

The Optimization of the MAGIC Telescope for Pulsar Observations

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Abstract

MAGIC, a new 17-meter class optical reflector at La Palma, Canary Islands, is the first Imaging Atmospheric Cerenkov Telescope (IACT) able to detect pulsed γ -rays from neutron stars. Simulations to predict detection times and expected γ -ray fluxes from millisecond and canonical pulsars are performed, using the standard operational mode of MAGIC. Spectral cutoffs due to the pair production of energetic γ -rays in strong magnetic fields are expected: The low surface magnetic field strengths of millisecond pulsars (relative to that of canonical pulsars), result in much larger spectral cutoffs for millisecond pulsars (i.e. above 100 GeV), whereas for canonical pulsars, this cutoff is expected to be ≤ 30 GeV, with the upper limit corresponding to the standard energy threshold of about 30 GeV for MAGIC. The relatively low spindown power associated with millisecond pulsars result in a low photon flux, which is however offset by their higher spectral cutoffs (around 100 GeV), where the MAGIC collection area is already relatively large. The result of this is that we still expect millisecond pulsar detection times of a few hours with MAGIC. Since most canonical pulsars are expected to show a cutoff at or well below 30 GeV, a new technique is proposed to detect γ -rays within the 2 to 10 GeV range, using the central ~ 0.4 degree region of the camera. Effective areas from 50 to 2500 m^2 are found, without interfering with the standard operational mode. When observing pulsars, timing parameter are required to perform a periodicity search. If contemporary radio parameters are not available, we have to extrapolate archival radio parameters to the observation

time in question. Such an extrapolation must then be accurate enough to avoid significant pulse smearing due to the intrinsic pulsar timing noise and glitches. A pulsar population study is done to estimate the effect of timing noise and glitches. But sometimes, the optical emission can be used to derive timing parameters. The use of a IACT camera central pixel to detect the optical pulsed emission is discussed. This method was tested with the HEGRA CT1 telescope and the application of this technique to MAGIC is also discussed.

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Chapter 1

Introduction

Astronomy is now performed over the entire range of the electromagnetic spectrum, from radio to γ -ray energies. From the so-called “New Astronomies”, which are performed outside the optical window, we learned that each spectral range provides specific information which cannot be obtained by other means. Gamma radiation represents the most energetic part of the electromagnetic spectrum (see Fig.2.2). Therefore it follows that it provides information about the most energetic processes and phenomena in the Universe [41, Gaisser, 1990].

Among these phenomena we concentrate on pulsars, a kind of rotating, magnetized, conducting star that forms a so-called unipolar inductor and which is capable of accelerating particles to reach relativistic energies, well above 1 TeV. The most compact and energetic objects emit γ -rays, not only neutron stars but also stellar and massive black holes, supernova explosions/remnants, and cosmic rays, via their interaction with matter and fields. In addition, it appears that most of the γ -ray sky is continuously changing. With γ -rays we see the most violent part of the Universe [95, Thompson et al., 1993].

For γ -rays, the particle description of electrodynamic radiation becomes more appropriate than the wave description, which works well for the less energetic radiation at radio or optical wavelengths. This helps us to understand the pen-

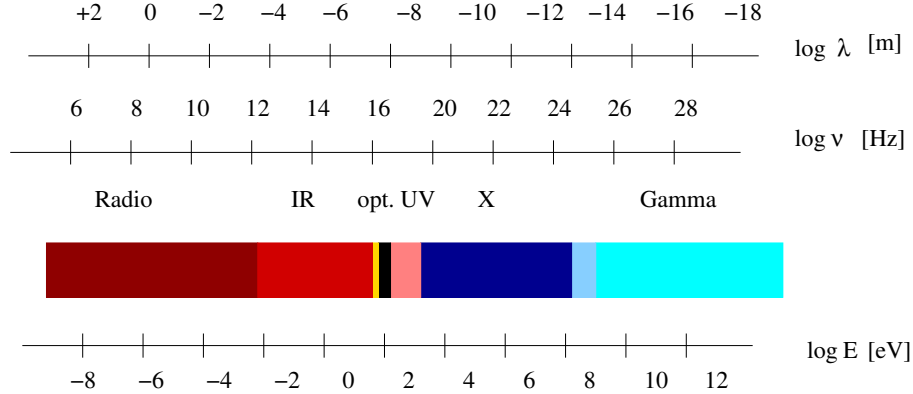


Figure 1.1: *The electromagnetic spectrum, from radio to γ -ray energies. The electromagnetic spectrum can be characterized either by its photon energy (measured in eV) or by its frequency (measured in Hz) or by its wavelength (measured in m).*

etration depth of high energy γ -rays: the wavelength of the radiation is short compared to the spacing of atoms in the material, hence the radiation mainly 'sees' the atom's components, comprising the nucleus and the electrons at comparatively large distances.

The energy band of γ -ray astronomy extends from typically $\gg 100$ keV to more than 1 TeV, and it can be separated into two broad domains. The first one is the domain of space-borne γ -ray astronomy, which ranges from 500 keV to about 10 GeV. γ -rays in this energy range cannot penetrate the Earth's atmosphere without being absorbed. These γ -rays can only be detected from space with satellite experiments or with high altitude balloon experiments. The second domain in which this thesis is developing observation techniques is that of ground-based γ -ray astronomy. This ground based technique operates at energies above 30 GeV for the Imaging Atmospheric Cerenkov Telescope (IACT thereafter) MAGIC telescope [66, Martinez et al., 2003], 50 GeV for HESS [51, Hoffman et al., 2001] and 100 GeV for VERITAS [99, Wakely et al., 2003] and CANGAROO [4, Asahara et al., 2003], the new generation of γ -ray telescopes.

At these energies primary γ -rays interact with atmospheric particles to produce Cerenkov emission, which can be detected with optical reflectors having fast photon counters in the focal plane. We will discuss, in this chapter, the basic physics processes to understand the production and propagation of high energy photons, along with the different techniques to detect such radiation. The final chapters (3,4 & 5) will cover detection techniques and observational constraints for pulsed radiation from such systems.

1.1 Gamma Ray Production Processes

Electromagnetic radiation may be either thermal or nonthermal. The first one emerges from a large population of electromagnetically interacting particles in equilibrium, with their mean energy characterized by a particle temperature, while the so-called nonthermal radiations do not require that the source particle spectrum follows a thermal population. Nonthermal processes are more typical sources of γ -rays, since under the hypothesis of Wien's law, $\lambda \cdot T = 0.2989 \text{ (cm K)}$, for a black body spectrum, a fireball with a temperature above $2 \times 10^9 \text{ K}$ would be required to produce a thermal γ -ray of 1 MeV. For instance, nuclear fusion inside the sun occurs at 10^7 K , corresponding to keV's in thermal energy. In comparison, γ -ray fireballs would have to be significantly hotter.

In general, all elementary particles which take part in an electromagnetic interaction may be sources of γ -rays, if accelerated in some way through an external field of any kind. Fig. 1.2 shows different processes for the creation of γ -rays.

1.1.1 The Acceleration of particles in Strong Electric Fields

For the purpose of this thesis, we will briefly discuss acceleration of charged particles in strong electric and magnetic fields: suppose that the environment in

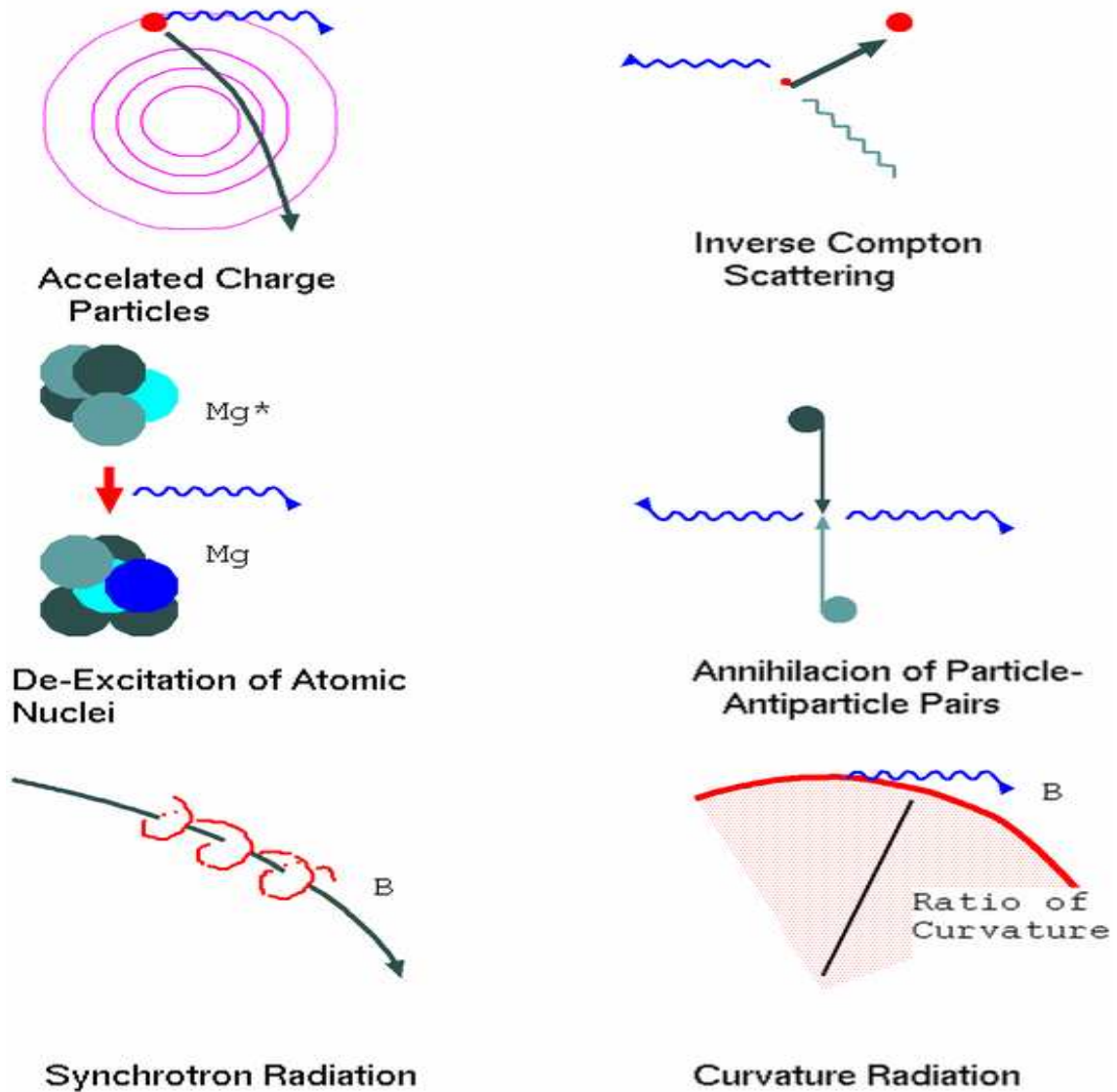


Figure 1.2: *Charged-particle acceleration results in photon emission. As an example for charged-particle acceleration the case of bremsstrahlung is illustrated (top left). Weak decays inside nuclei in excited nuclear states, which often decay through γ -ray emission (middle left). Likewise, the decay of unstable particles such as pions, and the annihilation of particle-antiparticle pairs constitute γ -ray source processes (middle right). Soft photons of energies lower than γ -ray energies may gain energy from electromagnetic-field interactions. The inverse-Compton process of energetic electrons or protons is illustrated (top right) as the most important example. On the bottom curvature and synchrotron radiation is plotted.*

a neutron star magnetosphere is such that an electric field $\mathbf{E}$ is generated through the pulsar dynamo effect (see Chapter 2). The parallel component of the electric field results in particle acceleration such that the particle energy increases at a rate of

$$\frac{dE}{dt} = m \cdot c^2 \frac{d\gamma}{dt} = 2e \frac{\mathbf{E} \cdot \mathbf{B}}{\|\mathbf{B}\|} \cdot v \quad (1.1)$$

where it was assumed that the particle of charge $2e$ and rest mass m moves at a speed of $v \sim c \cdot \cos\theta$, along the pulsar magnetic field $\mathbf{B}$.

Once ultrarelativistic energies can be reached, various mechanisms can cause the particle to loose its energy to γ -rays as discussed below.

1.1.2 Radiation of Charged Particles in Strong Electric or Magnetic Fields

The motion of a charged particle (e.g. electron) can be described as a charged current along the particle's trajectory. The particle's charge produces a Coulomb field; its motion thus corresponds to an electromagnetic field, varying as the particle moves. Any acceleration of the charged particle can be viewed as dynamic modification of this electromagnetic field, at the expense of the charged particle's total energy. Thus kinetic energy is transformed into energy of the electromagnetic field, which translates into the emission of γ -ray photons.

Two main radiation mechanisms are found to be important in strong magnetic fields to emit high energy emission, one is a superposition of curvature radiation of ultrarelativistic primary electrons, and the other a synchrotron radiation of secondary particles created via the magnetic absorption of photons.

(i) Curvature Radiation

Curvature radiation is produced when electrons or positrons are accelerated in vacuum gaps ($\mathbf{E} \cdot \mathbf{B} = 0$) along curved field lines, with pitch angle $\theta = 0$ given.

In neutron stars, primary electrons are injected along magnetic field lines into the magnetosphere from the outer rim of the polar cap. As electrons accelerate along the magnetic field lines they are simultaneously retarded by curvature radiation (CR) cooling, which represents the dominant cooling process:

$$(\dot{\gamma}_e)_{cr} = \frac{2}{3} c r_o \gamma_e^4 \rho^{-2}, \quad (1.2)$$

where γ_e is the electron Lorentz factor, ρ is the curvature radius, and r_o is the electron radius.

The energy of γ -rays depend on the Lorentz factor $\gamma = 1/\sqrt{1 - (v^2/c^2)}$ and the radius of curvature:

$$E_\gamma \propto \frac{\gamma^3}{\rho} \quad (1.3)$$

and therefore the γ -ray power

$$P_\gamma \propto \frac{\gamma^4}{\rho} \quad (1.4)$$

(ii) Synchrotron Radiation.

The motion of high energy charged particles in a magnetic field (with strength B) results in a synchrotron photon spectrum. This motion can be described then as the gyration of a particle around the field direction, with characteristic gyration frequency $\nu_g = eB/2\pi m_e$. The radiated energy originates from the velocity component, perpendicular to the magnetic field. The synchrotron spectrum peaks at a frequency:

$$\nu = \frac{3}{2} \gamma^2 \nu_g \sin\theta = \frac{3}{2} \gamma^2 \cdot \frac{eB}{2\pi m_e} \sin\theta \quad (1.5)$$

Applying to neutron stars, the electron/positron pairs radiate by synchrotron mechanism. The rate of synchrotron radiation cooling is:

$$(\dot{\gamma}_e)_{sr} = -\frac{2}{3} \frac{r_o^2}{m_e c} B_{ps}^2 \sin^2 \theta \gamma_e^2 \quad (1.6)$$

where θ is the pitch angle of the pairs (the angle between the particle trajectory and the direction of the magnetic field), B_{ps} the magnetic field in the pulsar surface and m_e the mass of the electron.

Both high magnetic field strengths and very energetic particles (large Lorentz factors) shift the radiation up in energy, and in particular, into the γ -ray region near the surface of neutron stars.

The γ -ray energy from synchrotron emission:

$$E_\gamma = 6.7 \text{ GeV} \left(\frac{B_{ps}}{10^{12}} \right) \left(\frac{E_e}{1 \text{ TeV}} \right)^2 \left(\frac{\sin \theta}{10^{-7}} \right) \quad (1.7)$$

(iii) Inverse Compton Scattering.

Up-scattering of photons of lower energy through collisions with energetic particles is called the *inverse Compton process*. When γ -rays collide with other particles such as electrons, it is known as the Compton scattering process. In *normal* Compton scattering, a γ -ray photon will collide with one of the many atomic electrons within some material and be scattered in the collision, transferring some of its energy to the electron. However the inverse energetic process may also apply and provide a process for γ -ray production: if photons of lower energy collide with energetic electrons, these photons may gain energy in the collisions, thus being upscattered in energy, e.g. from optical or X-rays to γ -rays. This process is important in regions of high photon densities.

When the energy of the photon in the center of momentum frame of reference is much less than $m_e c^2$, ($E_{ph} E_e \ll (m_e c^2)^2$), the typical energy of inverse-

Compton-scattered photons rises rapidly with energy for this limit, where Thompson scattering cross-section [60, Longair, 1981] can be used to describe the probability of scattering.

$$\sigma_c = \sigma_T \left(1 - \frac{2\gamma\epsilon}{m_e c^2}\right), \quad (1.8)$$

where σ_t is the Thompson scattering cross section $\sigma_t \approx 6.65 \times 10^{-25} \text{ cm}^2$.

The energy of scattered photons then:

$$E_\gamma \sim 1.3 \left(\frac{E_e}{TeV}\right)^2 \cdot \left(\frac{E_{ph}}{2 \times 10^{-4}(eV)}\right) \text{ GeV} \quad (1.9)$$

with an ambient photon-field typical energy E_{ph} and high energy electrons of energy E_e . This inverse-Compton scattering is also important in compact stars where an accretion disk is sufficiently hot to emit X-rays, and the compact object generates beams of charged particles in its vicinity.

In the case of pulsars, where $E_\gamma \approx E_e$ the correct scattering cross section for free electrons [55, Klein & Nishina, 1929] limit must be used:

$$\sigma_c = \frac{3}{4} \sigma_t \frac{mc^2}{\gamma h\nu} \left(\frac{1}{2} + \ln\left(\frac{2\gamma h\nu}{mc^2}\right)\right) \quad (1.10)$$

1.1.3 Charged Particles Bound in Strong Magnetic Fields

Strong magnetic fields can force electrons into orbits around their field lines, leading to quantized energy levels for allowed electron energies. Magnetic fields close to neutron stars can be strong enough for the steps between such allowed electron orbit levels to be in the regimen of tens of keV, the low γ -ray regime. Electron transitions from one allowed state to another will eject or absorb photons of this characteristic energy difference, producing the so-called cyclotron line radiation. These energy levels $E = h\nu_c$ may be obtained from the cyclotron

frequency

$$\nu_c = \frac{ZeB}{2\pi\gamma m_o} \quad (1.11)$$

for a particle with charge Ze and velocity v (hence with a Lorentz factor associated) in a magnetic field B . For field strength values around 10^{12} G as observed in strongly magnetized neutron stars, cyclotron lines fall in the X-ray regime, from $\hbar\omega \simeq 12 \text{ keV } B/(10^{12} \text{ G})$ [97, Trümper et al., 1978].

We have seen that, in general, for γ -ray production violent processes are at play. Observations of γ -rays, enable us to study such exceptional places in nature. We will discuss in the next section the different techniques used in high-energy astronomy.

1.2 Extensive Air Showers and Cerenkov Radiation

Three important facts govern the techniques that are used in high-energy γ -ray astronomy.

- The opaqueness of the Earth atmosphere for γ -rays. Although the density of the air is rather low ($1.293 \text{ kg}\cdot\text{m}^{-3}$ at ambient pressure) and although γ -rays are highly penetrating particles, the atmosphere is opaque for γ -rays because its integrated matter density amounts to $\sim 1000 \text{ g cm}^{-2}$. Since the mass-attenuation coefficient for air at 1 MeV is $\sim 0.00642 \text{ cm}^2\text{g}^{-1}$ the absorption probability for a 1 MeV γ -ray is $>99.8\%$. Thus, only a detector above the Earth's atmosphere, in a balloon or a satellite, can detect primary cosmic γ -rays (see Fig. 1.3 for different techniques of detection depending on the energy range).
- The fluxes of high-energy γ rays from astrophysical sources are quite low

and decrease rapidly with increasing energy. For instance, Vela, the brightest γ -ray source in the sky, has a flux above 1 GeV of 148.1×10^{-8} photon $cm^{-2} s^{-1}$ [58, Lamb & Macomb, 2000], and a differential spectrum that falls as $E^{-1.89}$ [48, Hermsen et al., 1981]. This implies that above some energy, a detector in a satellite will be too small to detect enough photons to be useful. A $\sim 1000 cm^2$ detector in a satellite would detect approximately five photon/minute from Vela above 1 GeV. The pursuit of γ -ray astronomy above 10 GeV energies must be done with much larger instruments than can be launched on a satellite. Earth-based detectors appear to provide a viable solution.

- The high intrinsic background of γ -ray telescopes. Charged particles such as cosmic ray protons [102, Wiebel-Sooth, 1998], electrons and He-nuclei are bent by the interstellar magnetic fields, so that they form an essentially isotropic background. The cosmic ray proton flux has been measured to be:

$$\frac{dF_p}{dE} = 10.91 \times 10^{-2} \cdot \left(\frac{E}{1TeV}\right)^{-2.75} \quad TeV^{-1}m^{-2}s^{-1}sr^{-1} \quad (1.12)$$

Background rejection is specially important at low energies when the usual method based in the **Hillas parameters** [100, Weekes et al., 198] becomes less efficient. This rejection technique will be discussed in more details in the following chapters, together with the consideration of different sources of background (Chapter 2).

Above 100 GeV, the γ -ray flux is so low that direct detection is impossible even with the proposed GLAST [35, Dingus et al., 1997]. The ground-based telescopes such as MAGIC, working from 1 GeV up to a few TeV are based on the detection of a **Extensive Air Shower** (from now on denoted as EAS) produced by a primary cosmic ray particle or photon. The cosmic radiation

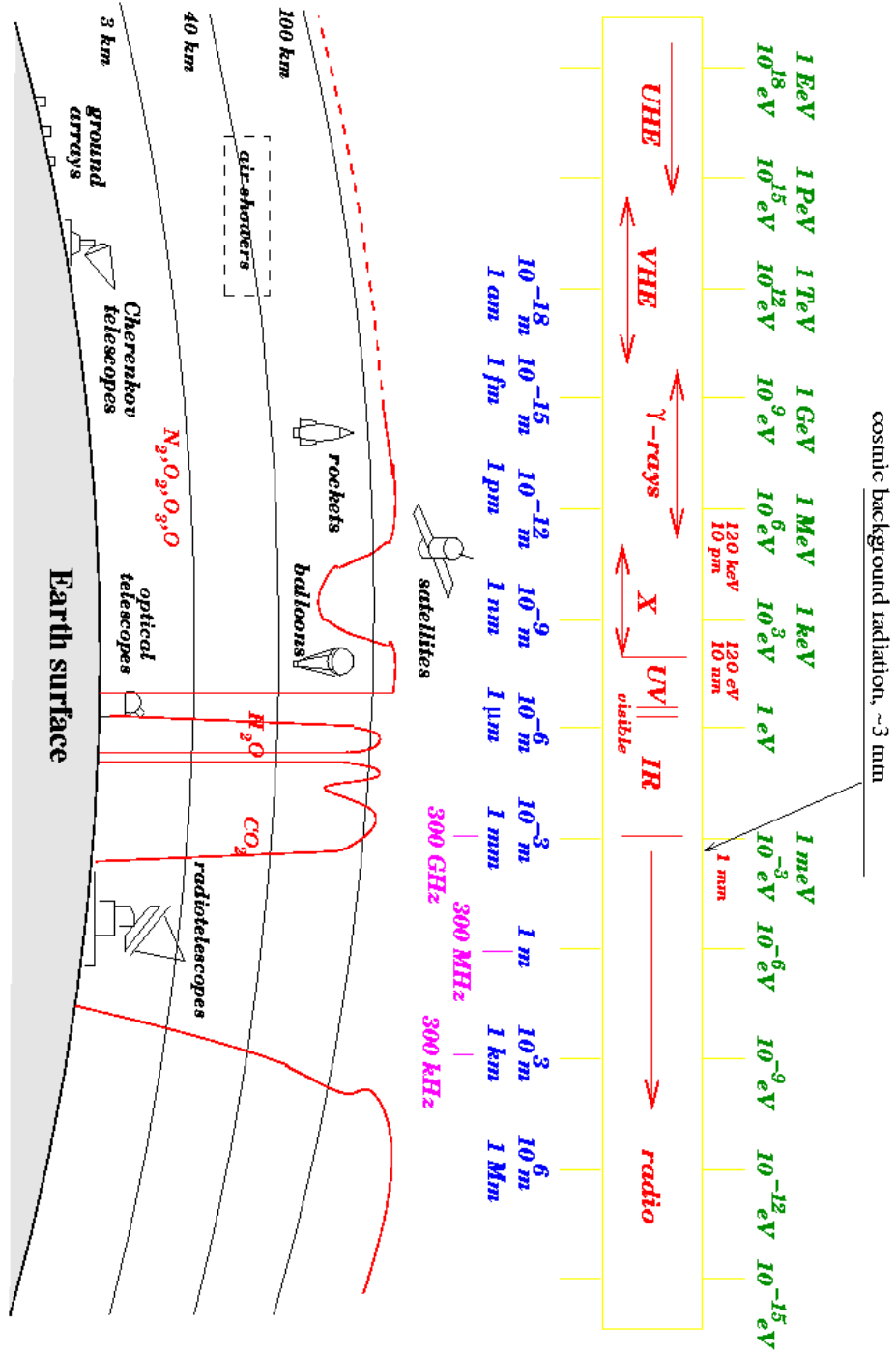


Figure 1.3: Different techniques used in high-energy astronomy to detect γ -rays, depending on the energy range. ([22, Cortina, 1997])

(CR), comprising atomic nuclei (96% H, 3%He, 1% C,N,O,Fe), γ -rays, electrons and positrons, neutrinos and other types of elementary particles, travels from the emission source until it reaches the Earth. On its way through outer space, some interactions with the intergalactic or interstellar medium can take place, as for example fragmentation of the nuclei, ionization, particle production and many more, so that secondary cosmic particles may eventually reach the Earth's atmosphere. When a cosmic ray particle enters the Earth's atmosphere, its super relativistic energy may give rise to a number of complicated processes after its first interaction with an atmospheric nucleus. This gives rise to a large number of *secondary particles* turning into EAS.

The primary particles in the EAS in the upper atmosphere are, mainly, nuclei, essentially and α -nuclei, and, in smaller numbers, heavier nuclei. The distribution of the incidence direction is basically isotropic as a result of the randomization of pitch angles by galactic and intergalactic magnetic fields. Amongst these high energy particles reaching the Earth, there is small flux of γ -rays. About 1 out 1000 air showers produced in the atmosphere are gamma initiated. Showers initiated by high energy electrons or γ -rays are more homogeneous with respect to secondary species compared to the ones initiated by a hadron. A high energy photon generates an electron-positron pair. This recursive process continues until the cascade energy is depleted.

The Cerenkov radiation detected by ground based telescopes involves radiation emitted by the medium under the action of the field of the particle moving in it as discussed below:

1.2.1 Properties of Cerenkov Radiation

Cerenkov light in the atmosphere is produced by charged particles traveling faster than the speed of the light in air.

This emission is governed by the refractive index (n), which is proportional

to the atmospheric density. The Cerenkov technique for detecting EAS then is based on the fact that electrons in EAS generate Cerenkov radiation if their energy exceeds a minimum threshold E_{min} . This threshold is 21 MeV at sea level and increases to 35 MeV at 7.5 kilometers above sea level. The reason for the change in E_{min} is the variation of the Cerenkov threshold velocity with index of refraction of the atmosphere

$$v = \frac{c}{n(H)}, \quad (1.13)$$

where c is the speed of the light in vacuum and n is the index of refraction at a given atmospheric height H . The simple geometric picture of this process (see Fig.(1.5)) is that, because the particle moves superluminally through the medium, a shock wave is created behind the particle and this results in a loss of energy by the particle.

From this figure it is understood that this radiation is only observed at a particular angle, Θ called the Cerenkov angle, with respect to the track of the particle. This angle represents the position in which waves from arbitrary points such as P1, P2 and P3 over the track AB are coherent and combine to form a plane wave front BC. This coherence takes place when the particle travels from A to B in the same time that the light travels from A to C. If the velocity of the particle is v or $\beta \cdot c$ and $\frac{c}{n}$ is the velocity of the Cerenkov light in the medium, then we can write the Cerenkov angle in the following way, taking only geometrical considerations into account;

$$\cos \Theta = \frac{\frac{c}{n(\lambda)} \cdot \Delta T}{\beta \cdot c \cdot \Delta T} \quad (1.14)$$

$$\boxed{\cos \Theta = \frac{1}{\beta \cdot n(\lambda)}} \quad (1.15)$$

From equation (1.15) we can derive the following:

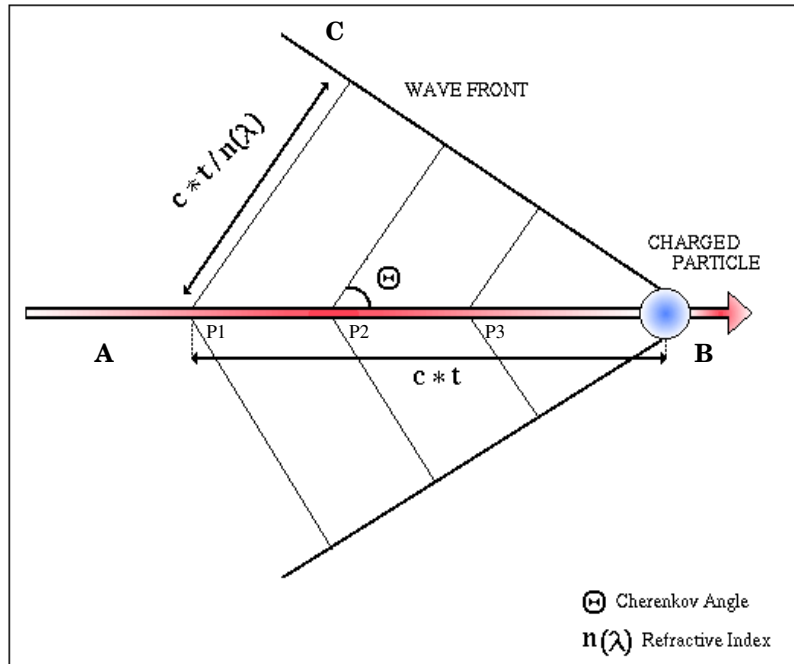


Figure 1.4: *Huygens construction to illustrate coherence of Cerenkov production along the particle track AB with the associated creation of a light from CB.*

- For a medium of a given refractive index n , there is a threshold velocity $\beta_{min} = \frac{1}{n}$, below which no radiation takes place. At this critical velocity the direction of radiation coincides with that of the particle. To see this we may write (2.12) in the following way:

$$\cos \Theta = \frac{1}{\beta \cdot n(\lambda)} = \frac{c_{medium}}{v_{part}} \quad (1.16)$$

so the process can be used in the construction of threshold detectors in which Cerenkov radiation is only emitted if the particle has velocity greater than $\frac{c}{n}$.

- For an ultra-relativistic particle, for which $\beta = 1$, there is a maximum angle of emission, given by,

$$\cos \Theta = \frac{1}{n} \quad (1.17)$$

- The radiation occurs in the visible and near visible regions of the spectrum, for which $n > 1$. A real medium is always dispersive, so actually radiation is restricted to those frequency bands for which $n(\omega) > \frac{1}{\beta}$. In the x-ray region $n(\omega)$ is always < 1 and radiation is forbidden. So emission in the x-ray region is impossible for n less than unity and Equation 1.17 cannot be satisfied.

There are two further conditions to be fulfilled to achieve coherence, in addition to that stated in the first point. First, the length l of the track of the particle in the medium should be large compared with the wavelength λ of the radiation in question. Otherwise diffraction effects will become dominant. Secondly, the velocity of the particle must be constant during its passage through the medium, or, to be more specific, the differences in the times for the particle to traverse successive distances λ should be small compared to the period $\frac{\lambda}{c}$ of the emitted light.

Therefore, EAS will emit a flash of Cerenkov light in a cone with opening angle Θ , typically of the order of 1° . This flash lasts for about 5 ns and yields about 50 photons $\cdot m^{-2}$ within 100 m around the axis for primary γ -rays of about 1 TeV. More accurately the number of photons emitted by a charged particle of charge Ze per unit path length and per unit energy interval, or equivalently, λ of the photons is equal to:

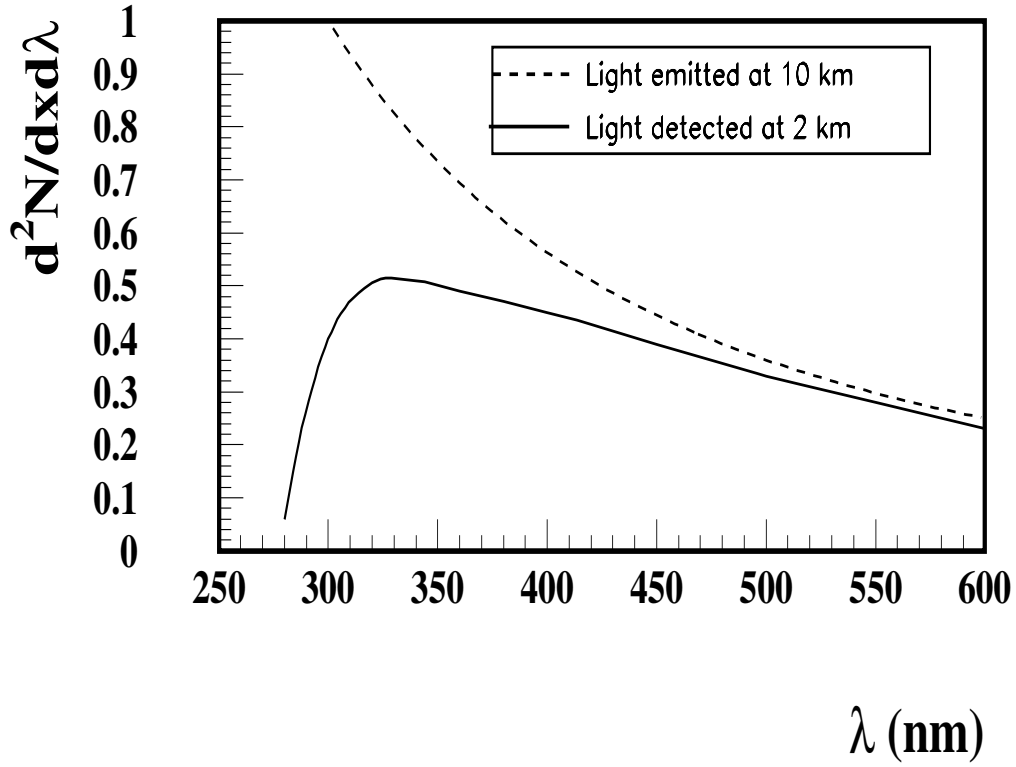


Figure 1.5: *Differential Cerenkov photon spectrum, in arbitrary units vr. wavelength in nm . Continuous line includes absorption by ozone, and scattering due to the Rayleigh and Mie effects. Figure provided by A. Moralejo [73, Moralejo, 2000], the dotted line shows the emitted light at 10 Km and the solid line the detected line at 2 Km.*

$$\frac{dN^2}{dxd\lambda} = \frac{2\pi\alpha Ze^2}{\lambda^2} \cdot \left(1 - \frac{1}{\beta^2 \cdot n^2(\lambda)}\right) \quad (1.18)$$

For the particular case of an electron moving along a track of length l within a spectral region defined by wavelengths λ_1 and λ_2 we will have:

$$N = 2\pi\alpha l \cdot \left(\frac{1}{\lambda_2} - \frac{1}{\lambda_1}\right) \cdot \left(1 - \frac{1}{\beta^2 \cdot n^2(\lambda)}\right) \quad (1.19)$$

where α is the fine structure constant $= \frac{e^2}{\hbar \cdot c} = \frac{1}{137}$ and n is the refraction index of the medium that is a function of the photon energy, or equivalently to λ . This means that the greater part of Cerenkov photons are emitted in the ultraviolet range, because $\frac{dN^2}{dx d\lambda} \propto \frac{1}{\lambda^2}$ and the spectrum has a peak at around 330 nm, when observed at 2 Km above the sea level as shown in Fig.1.5.

The very diffuse nature of the flash together with the distance to the event (about 10 km) requires large optical mirrors for the collection of a significant fraction of the light from the flash. In order to detect this flash it has first to be distinguished from the background of normal light in the atmosphere. The maximum emission of the light from the night sky lies towards the red end of the spectrum, enabling one to choose the photomultipliers (PMs) response of the IACT instrument accordingly. The discrimination against showers initiated by other particles like hadrons, which are created as secondary particles from cosmic rays, is crucial for the IACT to be successful. Two approaches are used: the wave-front sampling technique, where one tries to measure the propagation of the Cerenkov wave front for discrimination between hadronic and electronic showers, and the more successful imaging technique. The later technique will be discussed in more detail below:

1.2.2 The Imaging Technique

The imaging technique exploits the differences between electromagnetic and hadronic cascades:

- **Electromagnetic Cascades (γ -ray & electron primaries)**

At the energies involved in the process that takes place in the upper atmo-

sphere, when a γ -ray plunges into the atmosphere, pair production is the dominant effect for EAS. The process of pair production takes place in the field of an atmospheric nucleus or electron in order to conserve momentum. A high energy photon in matter, of at least 1.022 MeV, will convert into an electron and positron pair. If the energy is sufficient, the resulting electron positron will produce a secondary γ -ray in the field of an atmosphere nucleus before the pair can loose its energy due to ionizing interactions. This process is called bremsstrahlung. This γ -ray, if its energy is still higher than 1.022 MeV, can then produce another electron positron pair that can undergo further bremsstrahlung interactions. The result of this recursive process is a cascade of photons, electrons and positrons, which continues traveling in the original direction of the primary γ -ray and sharing its total energy in the cascade.

- **Hadronic Cascades (proton, He nucleus and heavier primary elements)**

These particles, most of which are protons, cause nuclear disintegration at the top of the atmosphere, with the emission of further particles which in turn themselves produce further spallation products. We could write this interactions in the following way:

$$\text{Cosmic Ray (CR)} + \text{Atmospheric Nuclei (AN)} \Rightarrow \text{CR}' + \text{AN}' + n\pi^{\pm} + m\pi^0 + \text{other mesons},$$

where CR' is a fragment of the original cosmic ray that can keep producing more interaction with atmospheric nuclei. If the original cosmic ray has enough energy, some of its fragments could reach the ground. AN' are fragments of atmospheric nuclei in a high energy state. Among the products of the nuclear disintegrations are the π -mesons, which can decay in the following sub-products,

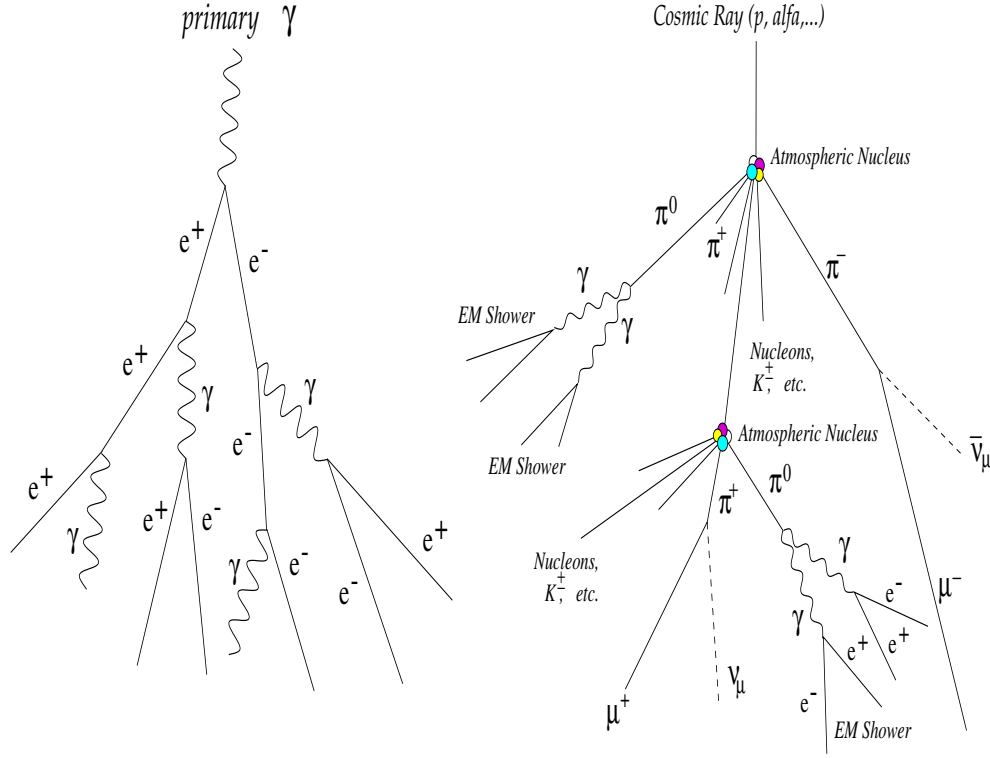


Figure 1.6: *Geometric model of emission of Cerenkov radiation for γ -ray and hadron shower. Provided by A. Moralejo [73, Moralejo, 2000]*

$$\begin{aligned}
 \pi^0 &\rightarrow 2 \gamma & \tau &= 1.8 * 10^{-16} sec \\
 \pi^\pm &\rightarrow \mu^\pm + \nu_\mu & \tau &= 2.5 * 10^{-8} sec \\
 \mu^\pm &\rightarrow e^\pm + \nu_e + \nu_\mu & \tau &= 2.2 * 10^{-6} sec
 \end{aligned}$$

The decay of the π^0 in two high energy γ -rays results in the development of electro-magnetic component of the shower because from here on only electrons and γ -rays will be produced, having a component made up by electrons and γ -rays by the processes of bremsstrahlung and pair production as discussed previously. The process by which this sub-cascade develops, is the same as the one described for the pure electro-magnetic showers. Fig. 1.6 shows a schematic view of the development of a hadronic and electromagnetic cascade. It is important to notice that hadronic showers

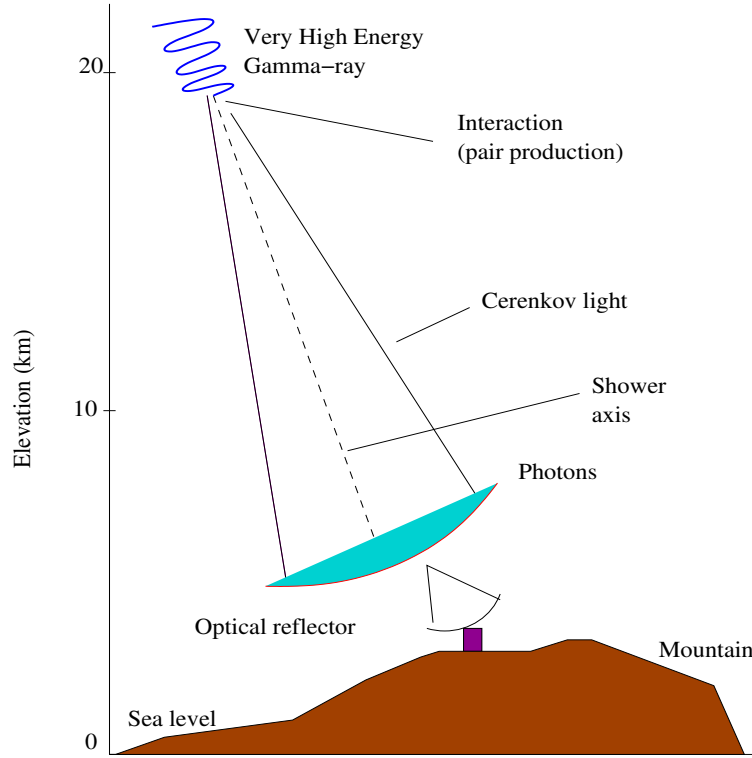


Figure 1.7: *Very high energy γ -ray impacting with the atmosphere and developing a showers of particles. The Cerenkov front of light keeps the direction of the shower axis, where the telescope is pointing at. The light is collected with a optical reflector and driven to the camera located at the focal plane.*

are more spread out than electron-photon cascades.

The objective of the atmospheric Cerenkov telescopes is to obtain an image of the Cerenkov light from the shower in the detector. This image represents a geometrical projection of the shower into a detector, provides that the angular surface brightness of Cerenkov light is preserved with a high resolution camera to discriminate between the light of the night sky (LONS), γ -ray initiated showers and hadron initiated showers [101, Weekes & Turver, 1997].

The γ photons emitted by a γ -ray source arriving at the Earth's atmosphere, see Figure 1.7, keep their original direction of emission. This means that the shower axis will coincide with the telescope's optical axis, if the later is pointed

towards the source of γ -rays. In contrast, the charged hadronic component of the cosmic rays does not keep track of the original direction of emission due to the galactic magnetic deflection as discussed previously. The arrival of this background component is therefore isotropic, and one can expect a difference between hadronic and γ -ray showers in the front plane of the detector.

Consider showers for which the core position coincides with the telescope, whereas the shower axis is parallel to the optical axis of the telescope. Such showers produce circularly symmetric image. As the point of impact of the shower on the ground is displaced away from the detector, images become progressively more elongated and comet shaped. The position of the maximum light intensity no longer corresponds to the core location. The true direction of the shower moves away from the point of maximum light intensity toward the head of the image, but always lying on the major axis of the elliptical image.

By measuring the shape of each shower image, and selecting only those events which are γ -ray-like in appearance, nearly all the cosmic ray contamination can be removed, resulting in a much improved ability to detect an excess number of counts from the source direction.

The process of γ -hadron separation begins with some form of parameterization of shower images on either a second moments approach or else on some form of semi-analytical fitting of shower images [59, Le Bohec et al., 1998]. Given the essentially elliptical nature of the shower images it was natural that the parameterization of images was originally performed in terms of a moment analysis of the recorded pixel signal amplitudes. Moments are calculated from the ADC counts in each pixel, together with the particular pixel coordinate, thus image parameters are reconstructed from the pixel information. With finer pixels the shower images will contain more information which could be used for further improvements in γ -hadron separation.

The ellipse parameters can be classified as *shape parameters* which characterize the size of the image and *orientation parameters*. In figure (1.8) an ellipse is

drawn to represent a shower image. From this single figure we could take out the shape parameters:

- **Size:** The total integrated light content of the shower. Sometimes another parameter similar to size is used, and that is **Conc**, which represents the degree of light concentration as determined from the ratio of the five largest pixel signals to the sum of all signals.
- **Length:** The r.m.s. spread of light along the major axis of the image. It carries information of the longitudinal development of the shower.
- **Width:** The r.m.s. spread of light along the minor axis of the image. This parameter carries information of the lateral development of the shower.

If we place an image like this onto a camera some more quantities are needed to describe an image. These quantities are the orientation parameters:

- **Distance:** The distance from the centroid of the image to the center of the field of view of the camera, which gives information of the impact parameter of the shower with respect to the Cerenkov telescope (CT). From another point of view if we take a line to join the shower maximum and the telescope center, this is equivalent to the angle between that line and the shower axis.
- **Miss:** The perpendicular distance between the major axis of the image and the center of the field of view of the camera. This is a measure of the shower orientation.
- **Azwidth:** The r.m.s. spread of light perpendicular to the line connecting the centroid of the image to the center of the field of view. This represents the projection of width along a line which is perpendicular to a line joining the center of the camera and the center of the image and which contains the centroid. This is a measure of both the shape and orientation of the image.

- **Alpha:** Is the angle between the major axis of the image and the radius drawn from the center of the camera to the center of the image. This parameter is related to the angle between the shower axis and the axis of the telescope.

Other parameters have been proposed from time to time for the analysis of the images but the ones listed above have been proved to be more efficient.

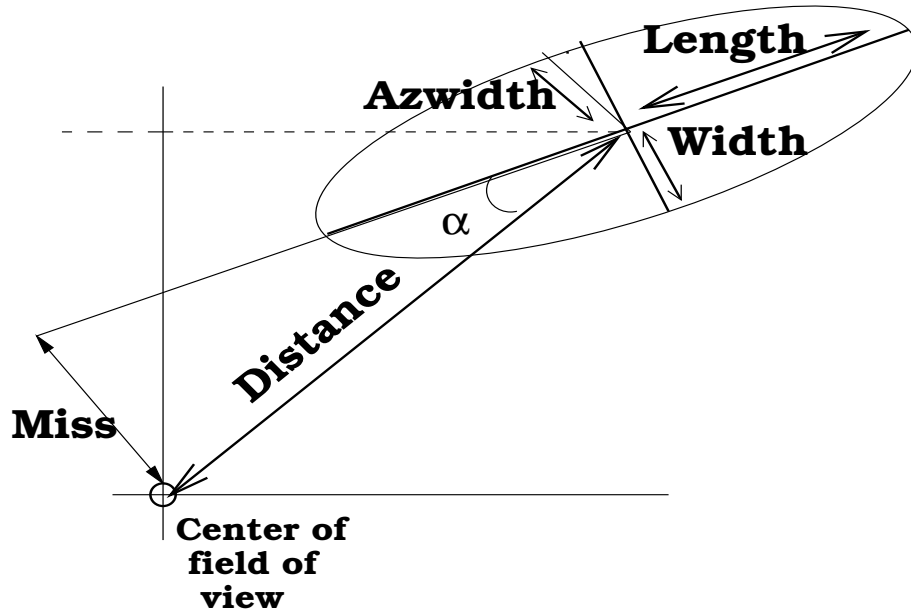


Figure 1.8: *Image parameters defined over an ellipse in the plane of the camera. Those parameters help us to discern between γ and hadron primary showers.*

1.3 Air-Shower Cerenkov Telescopes

Detectors based on the atmospheric Cerenkov technique consist of one or more mirrors that concentrate the Cerenkov photons onto fast optical detectors. Photomultiplier tubes (PMTs) placed in the focal plane are generally used to detect the Cerenkov photons.

Light from different heights is focused over different point in the focal plane, showing in this way onto the camera, an image of the longitudinal development of the cascade (number of particles emitting Cerenkov light versus height). This technique is known as imaging as it was explained.

Images reconstructed with an angular resolution of 0.25° have shown to be successful in detection TeV γ -ray showers (HEGRA [6, Barrio et al., 1998]), but an improved angular resolution of 0.1° [20, Cortina et al., 2003] is expected to field better results in term of γ -hadron separation.

Imaging Air Cerenkov Telescopes require clear moonless nights for optimal sensitivity. The energy threshold, between 250 GeV and 1 TeV (except MAGIC with a threshold of ~ 10 GeV, assuming a phase II with a Hybrid Photo Detectors camera), is determined by the minimum number of Cerenkov photons collected to distinguish signal above the LONS fluctuation. Therefore, it will be dependent on the night sky background flux ϕ , the mirror surface area A , detection efficiency ϵ , the solid angle on the sky viewed by the mirror Ω , and on the camera integration time τ , which will be fixed with the typical duration of a pulse of Cerenkov light [60, Longair et al., 1981].

$$E_{threshold} \propto \sqrt{\frac{\phi \Omega \tau}{\epsilon A}} \quad (1.20)$$

Although experiments with a system of several telescopes have larger mirror area considering the sum of the areas of individual telescopes, the trigger level is done by one single telescope and therefore the energy threshold is set by the mirror size of one telescope. In such a way, MAGIC with a mirror area larger than any single telescope currently working has the lowest threshold at the moment, decreasing it to a few GeV.

In Table 1.1 an overview of the existing and of some proposed instruments is given. The MAGIC telescope will be described in the next section.

Experiment	Situation hei., lat., long. (m, °, °)	Mount × dish	Dish area (m ²)	Pixel size (°)	Vision field (°)	Thresh (GeV)
Whipple-10m	Mt. Hopkins, Arizona 2300 32N 111W	1 × 1	72	0.25	3	250
MAGIC	La Palma, España 2200 28N 17W	1 × 1	230	0.1	4	10
ECO-1000	La Palma, España 2200 28N 17W	1 × 1	1000	0.1	4	5
Veritas	Mt. Hopkins, Arizona 2300 32N 111W	7 × 1	100	0.15	3.5	100
HESS	Khomas Highland, Namibia 1800 23S 15E	5 × 1	108	0.16	5	100
CAT-imager	Themis, Francia 1650 42N 2E	1 × 1	17	0.12	3	250
Durham-Mk VI	Narrabri, Australia 250 31S 110E	1 × 3	42	0.25	4	300
HEGRA-CT system	La Palma, España 2200 28N 18W	5 × 1	8.5	0.25	3.25	500
HEGRA-CT1	La Palma, España 2200 28N 18W	1 × 1	10	0.25	3	700
Telescope Array	Dugway, Utah 1600 40N 113W	3 × 1	6	0.25	4.5	600
CANGAROO	Woomera, Australia 0 31S 136E	1 × 1	11	0.19	3	1000
CANGAROO II	Woomera, Australia 0 31S 136E	2 × 1	10	0.19	4	200
CAO-GT-48	Crimea, Ucrania 600 45N 34E	2 × 3	4.5	0.4	3	1000
TACTIC	Mt. Abu, India 1300 25N 73E	1 × 1	9.5	0.31	3.2	1000
Lebedev-SHALON	Tien Shan, Kazakhstan 3338 43N 77E	1 × 1	10	0.6	8	1000

Table 1.1: *An overview of very high-energy γ -ray experiments, [50, Hoffman et al., 1999].*

1.4 The HEGRA CT1 and MAGIC Telescopes

The most sensitive detector above 10 GeV is MAGIC (**M**ajor **A**tmospheric **G**amma **I**maging **Č**erenkov Telescope [66, Martinez et al., 2003]), which is a 17 m diameter imaging detector, and which is expected to be much more sensitive than other IACT for the same observation time.

MAGIC is located next to the former HEGRA experiment, at the Roque de los Muchachos site, which is situated on the Canarias island of La Palma, a volcanic island off the African coast at 28°N and 17°W. The site condition for optical observations is excellent, and it is run by the IAC (Instituto Astrofísico de Canarias). The altitude above the sea level is from 2200 to 2500 m.

1.4.1 The HEGRA Detector

Originally, the experiment [57, Kranich, 2001] was built as a small scintillator array in 1988. The HEGRA collaboration was created by seven institutes: University of Hamburg, Max-Planck-Institut für Kernphysik in Heidelberg, University of Kiel, Universidad Complutense de Madrid, Max-Planck-Institut für Physik in Munich, University of Wuppertal and the Yerevan Physics Institute. In its 1997 setup, the experiment consisted of 17 Geiger counters, 224 scintillator counters, 77 wide angle Čerenkov counters and 6 Čerenkov telescopes. Of the 6 telescopes, the first one, CT1, has been used for certain studies in this thesis. Figure 1.9 shows the HEGRA experiment in 1997. In 2003, only CT1 is still working being part of the MAGIC collaboration, and the rest of HEGRA CT-system experiment has been dismantled. The later HEGRA system of telescopes has been one of the most successful experiment at very high energies [47, Heinzlmann et al., 2003].

The first Čerenkov telescope CT1 (see Fig.1.10) was installed in 1992. It comprised an equatorial mount, a 10 m² segmented all-aluminum mirrors and a fast imaging 37 high resolution 127 pixel, with $\sim 3^\circ$ field of view (FOV). CT1 is operated as a stand-alone telescope with a threshold of ~ 800 GeV. More detailed

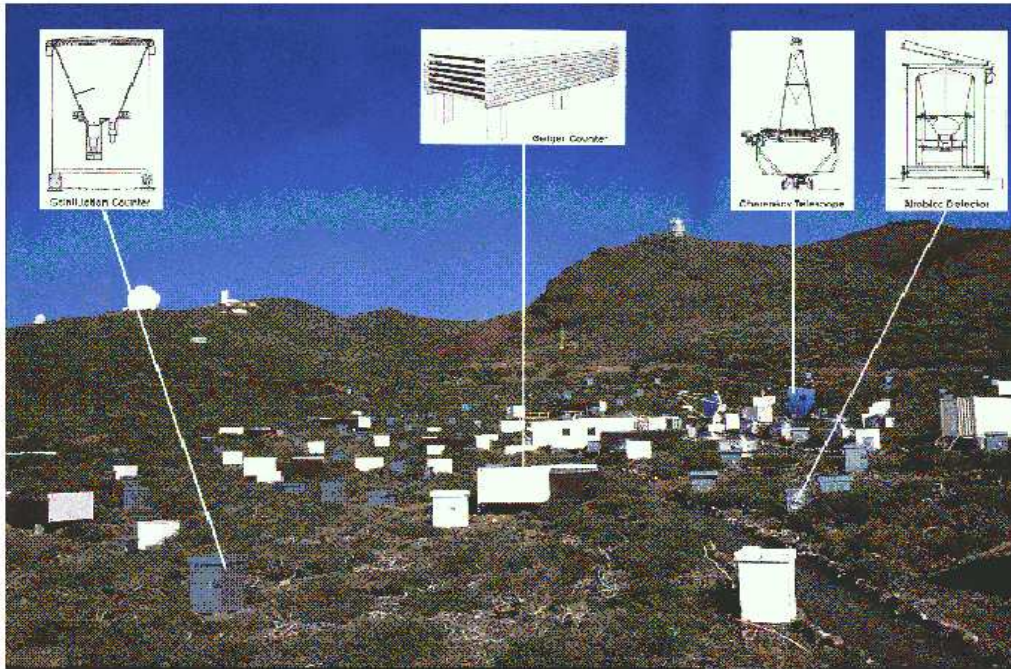


Figure 1.9: *The HEGRA experiment as it was in 1997. At the present moment, only one of the Cerenkov telescope CT1 remains, while the rest of the experiment has been dismantled (Courtesy of the HEGRA Collaboration)*

information about the CT1 camera and the readout electronics is described in by Rauterberg et al. [83, Rauterberg et al. 1995].

1.4.2 The MAGIC Telescope

MAGIC represents one of the major next generation telescopes. The MAGIC collaboration has developed novel innovations and new techniques to make it a superior instrument for VHE γ -ray physics. MAGIC is currently in its commissioning phase.

The MAGIC telescope project is a continuation of work that started before 1987 with the HEGRA detector. The first "letter of Intent" was presented



Figure 1.10: *The CT1 "prototype" telescope, part of the MAGIC collaboration at the moment (Courtesy of the HEGRA Collaboration).*

in Madrid in July 1998, and the official inauguration was on October 10th, 2003, although the first light was recorded on the 8th March (2003). The large MAGIC collaboration consists of the following institutes and universities: Institut de Física d'Altes Energies of Barcelona, Universitat Autònoma de Barcelona, Crimean Astrophysical Observatory, University of California, Davis, USA, Institute for Particle Physics, Swiss Federal Institute of Technology (ETH) Zurich, Division of Experimental Physics, University of Lodz, Universidad Complutense de Madrid, Institute for Nuclear Research, Russian Academy of Science, Moscow, Max-Planck-Institut für Physik, München, Dipartimento di Fisica, Università di Padova, Space Research Unit, Potchefstroom University, Detektorphysik und Elektronik, Fachbereich Physik, Universität-GH Siegen, Dipartimento di Fisica,

Università di Siena, Tuorla Observatory, Pikkio, Finland, Dipartimento di Fisica dell'Università di Udine/Trieste, Universität Würzburg and Yerevan Physics Institute, Cosmic Ray Division, Yerevan.

The MAGIC telescope [66, Martinez et al., 2003] was designed in 1998 with the main goal of being the IACT with the lowest γ -energy threshold possible.

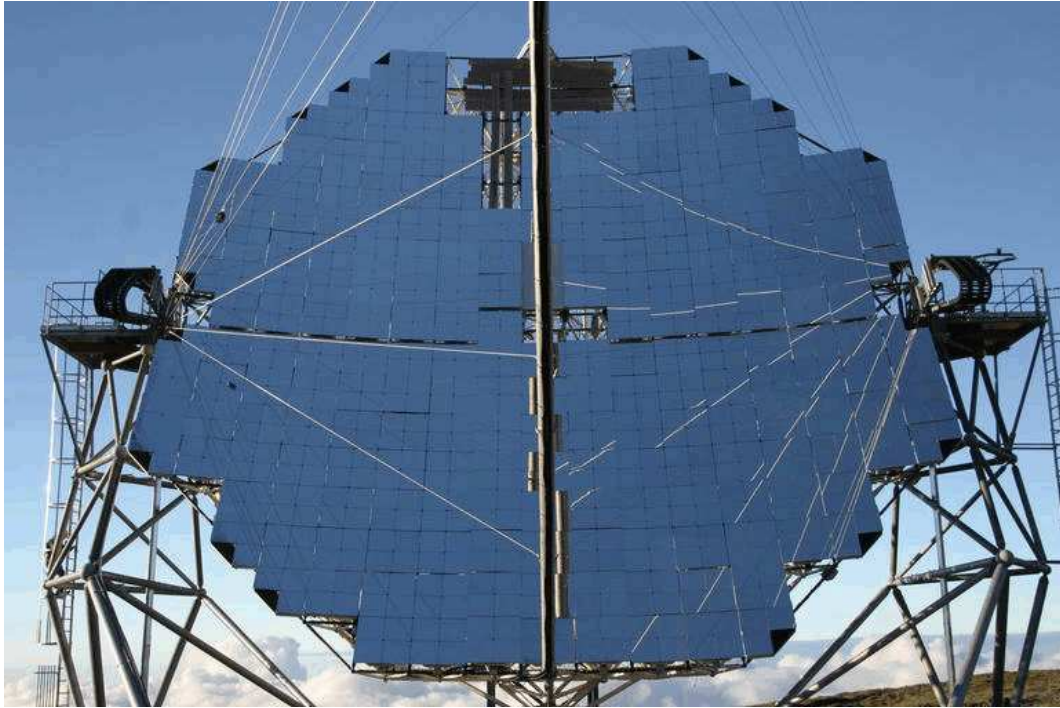


Figure 1.11: *The MAGIC telescope with all the mirrors installed*

The main motivation to achieve this low threshold is based on the sky map revealed by EGRET, the high energy γ -ray detector aboard the Compton Gamma-Ray Observatory (CGRO, [95, Thompson et al., 1993]). There is a well populated sky-map of sources detected by EGRET up to 10 GeV (most of them still unidentified due to poor angular resolution). Fig. 1.12 shows a sky map based in the third EGRET catalogue (http://coss.gsfc.nasa.gov/egret/3rd_EGRET_Cat.html). On the contrary, Fig.1.13 represents just a handful of sources observed by the existing IACTs above 300 GeV.

The catalogue of GeV EGRET sources constructed by Lamb & Macomb [58]

represents an important link between the bulk of EGRET sources (mostly seen above 100 MeV) and MAGIC sources (to be seen above 30 GeV).

THIRD EGRET CATALOGUE OF GAMMA-RAY POINT SOURCES

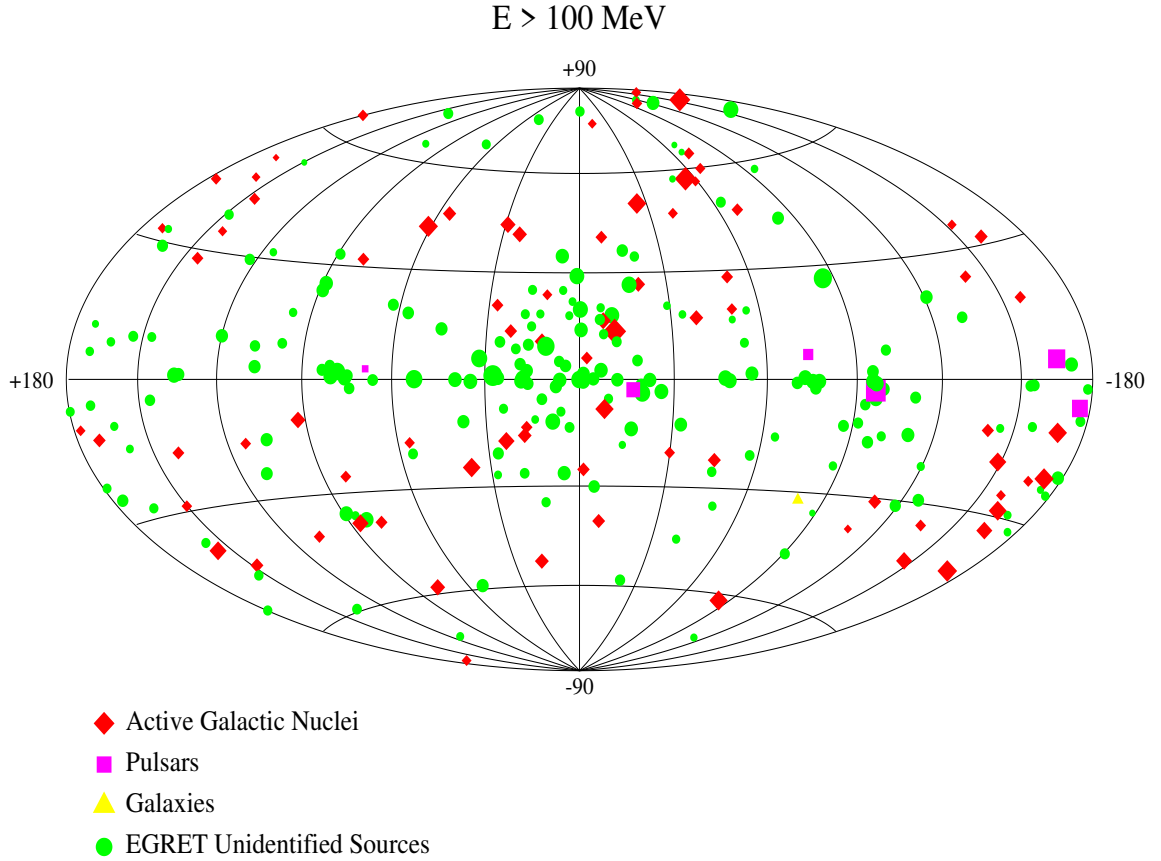


Figure 1.12: *EGRET* sources, 3rd catalogue up to $\sim 10 \text{ GeV}$.

The most critical parameters of the MAGIC telescope are the following:

- A 17 m diameter parabolic carbon-fiber frame with $f/D=1$, able to reposition in any direction in less than 30 seconds.
- A tessellated reflector made of 956 half squared meter diamond-turned aluminum mirrors.

VHE Gamma-Ray Sources

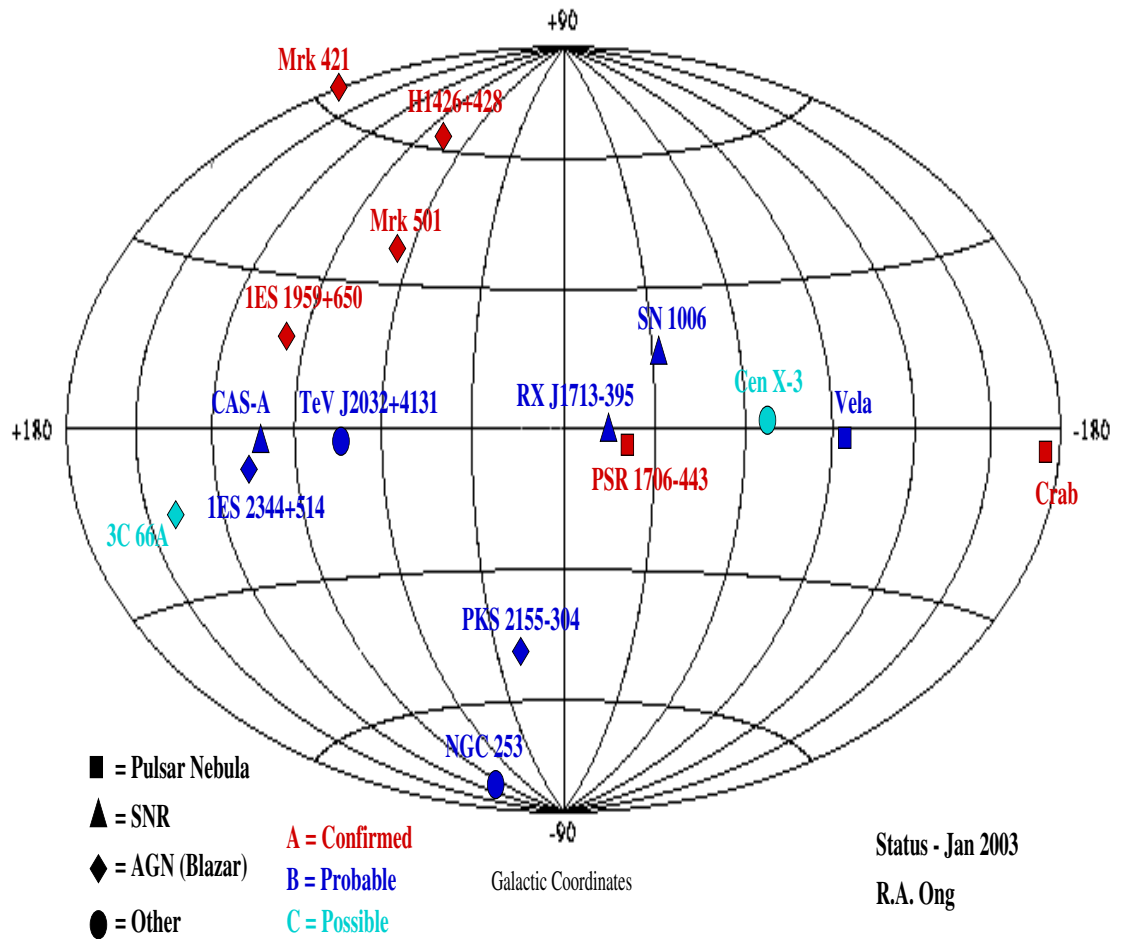


Figure 1.13: *VHE* sources detected with different ground based telescopes for energy higher than 300 GeV [76, Ong, 1998].

- A novel Active Mirror Control (AMC) able to correct the mirror pointing online for dish deformations.

- A camera made of 577 good quantum efficiency, fast photomultipliers with hemispherical photocathode to allow for light double-crossing and a special wavelength-shifting coating to provide red extended sensitivity and allow for light-trapping.
- Analogue signals are transmitted from the camera to the control house via optical fibers; only the amplifiers and laser diode modulators for transmission are inside the camera housing.
- A multilevel trigger and a 300 MHz FADC system for pulse digitization.

MAGIC will have the best light collection that has been attempted so far. As a result of all these improvements, it will be more sensitive to electromagnetic showers at lower energies. In the next chapter the main features of pulsar spectrum, characterized by a cutoff at energies between 10-300 GeV are discussed. Thus, MAGIC then is essential in understanding the physics involved in pulsar magnetosphere.

Others research targets will be:

- **Active galactic nuclei:** Recent results indicate that most if not all galaxies (including our own milky way) have an active center, in which a super-massive black hole is building up. Some of them (Mrk 421, Mrk 501) have been observed to be active in the VHE γ region, with occasional outbreaks and even with quasi-periodic fluctuations. The preferred theory explains the VHE gammas as products of high acceleration fields (shock waves) in the jet that bundles charged particles along two directions at 180 degrees to each other. We currently believe that the VHE γ -rays are produced within the jets, close to the black hole.

The origin of the jets is not yet understood. Models relate the jet directions (seemingly constant over millions of years) to the spin axis (axis of rotation) of the black hole. Understanding more about these objects and

the acceleration mechanisms both in the vicinity of the black holes and in intergalactic space is a task in which IACTs have an important role to play. MAGIC, in particular, with its emphasis on optimal light collection will be able to probe more deeply into the earlier part of the developing universe: the lower the energy threshold, the larger the observable red-shifts [5, Bastieri et al., 2003].

- **Supernova remanent:**

In the wake of a certain class of supernova explosions, the so-called SN of types II and Ia, gas clouds expand and a very dense core develops; the core may be a spinning neutron star or a black hole. In the example of the Crab nebula, the neutron star is observed as a pulsar, because it rotates at 30 cycles and 'pulses' in the X-ray domain; it is also observed at optical and UV wavelengths.

A rather constant radiation at higher energy, in the TeV range, has also been observed by IACTs. Supernova remnants may be VHE γ sources of different types. According to the standard model of cosmic ray origin, the shell type supernova remnants (radiating from the expanding cloud) are sites of acceleration of nuclei to very high energies: if so, they not only are the main accelerators of charged cosmic rays, but should also copiously produce γ -rays.

SN remnants of the plerion type, instead, are expected to radiate from the core. A systematic high-sensitivity scan of candidates, most of them lying in our own galaxy, is yet to be done.

- **Sources found at lower energies but not yet identified:**

All-sky surveys of wide-angle searching experiments in satellites have discovered a large number of lower energy γ -ray emitters. The angular resolution of these detectors is limited, as we discussed before, so that for more

than half of these sources, it is not yet possible to relate them to known sources observed at different wavelengths. The (third) 1999 catalogue of sources established by the EGRET detector is a well-known book of astrophysics riddles. Many of the unidentified populate the galactic equator, hence can be expected to be in our own galaxy.

- **Gamma Ray Bursts:**

Discovered 30 years ago, these Gamma Ray Bursts have been objects for research and speculation ever since. One theory, the fireball or hypernova model, posits that they are indicative of extremely violent explosions releasing in excess of 10^{51} ergs (or 10^{44} J), and creating violent shockwaves as the materials flowing out from the explosion at different velocities collide [31, de Paolis, 2000].

Today, a few thousand γ -ray bursts have been carefully charted, mostly by the BATSE satellite experiment, now removed from orbit. These objects cover the entire sky, seem spatially uncorrelated, many of them have large red-shifts, i.e. we observe them at billions of years in the past, in the period of active star formation.

Observations in the very high-energy γ domain are not available, so far, but are expected to help clarifying these phenomena.

- **Other contributions to cosmology and fundamental physics:**

Observations of VHE gammas, if done systematically, will also allow to formulate constraints on stellar formation in the early universe, by measuring the extragalactic infrared radiation field. They will further allow searches for the stable lightest supersymmetric particle, expected (if it exists) to annihilate with its own self-conjugate antiparticle into photons in areas of high gravitational field, e.g. in the vicinity of the black holes of galactic centers (including our own).

Quantum gravity effects might become apparent if subtle time differences can be detected in the arrival of gammas from a given source, at different wavelengths. If they occur in nature, the MAGIC detector has the capability to record such phenomena [8, Blanch, 2003].

1.5 Motivation

MAGIC is the stand alone telescope with the best light collection that has been attempted so far. As a result of improvements discussed above, it is more sensitive to electromagnetic showers at lower energies. Figure 1.14 shows the comparison with other experiments. It is clear that the MAGIC telescope is the most sensitive instrument at GeV energies at present. The initial sensitivity given in the proposal is plotted (in blue in the figure) [7, Bigongiari et al., 2003]. The sensitivity of the future experiment ECO-1000 is also represented.

According to the results presented by EGRET, we intend to observe a very interesting range of energy in the spectrum. Our purpose is to improve the MAGIC sensitivity in the 1 to 30 GeV energy range. The main motivation for having a very large collection area at 2-10 GeV is to detect new (weak) pulsed/variable sources, since most canonical pulsars' γ -ray spectra cut off above 10 GeV. They are very bright below 10 GeV, but very faint above 10 GeV. The spectra below 10 GeV are also very hard, so you have an ideal window in the 2-10 GeV range for new pulsar discoveries using a new technique.

Expected fluxes and required observation times for canonical and millisecond pulsars must be estimated to determine the efficiency of detection of pulsars with MAGIC. An attempt has been done, not only to improve the technology already installed in MAGIC, but to open new possibilities of detection at very low energy, although without energy resolution.

In this thesis the possibility of using an IACT as an optical telescope as well as γ -ray telescope is discussed, together with several tests made to detect the

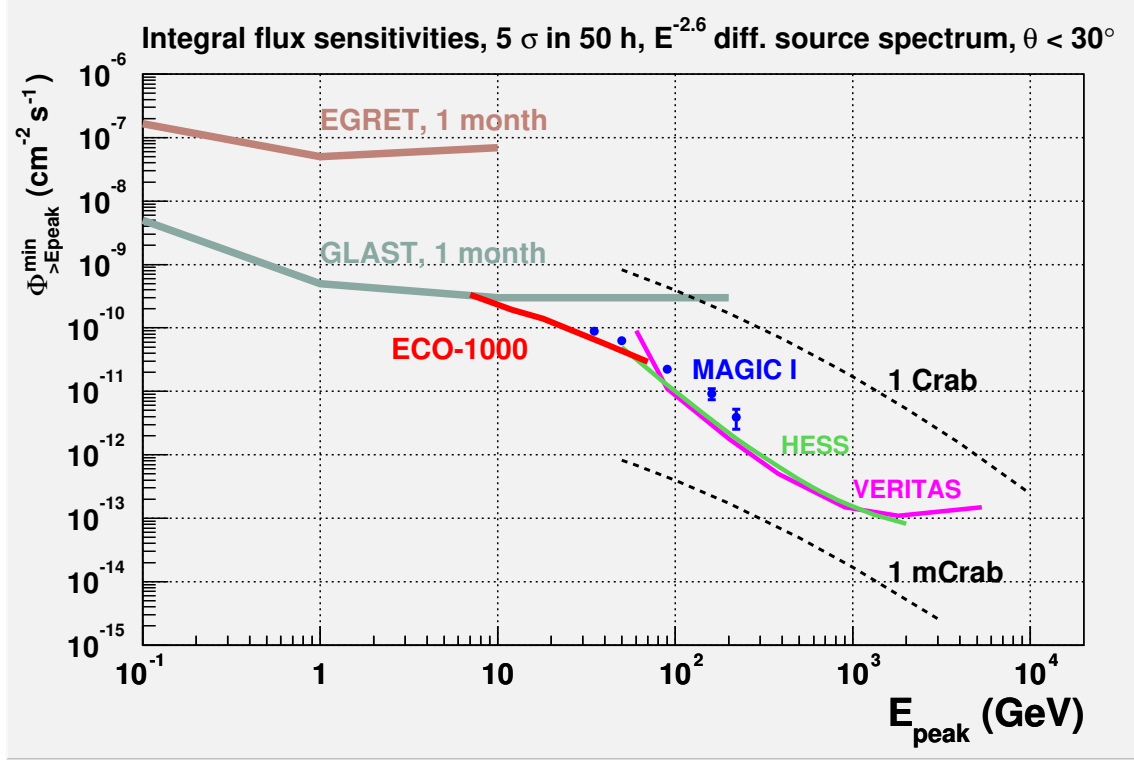


Figure 1.14: *Sensitivity of HE and VHE γ -ray Observatories.*

optical Crab Pulse. The method was tested in CT1 to be applied in the next generation of Cerenkov telescope, MAGIC. The recording of images in a Cerenkov telescope has quite different constraints from what an optical telescope requires. We will use the central pixel of the CT1 camera. The angular resolution is poor compared with a optical telescope, but the extremely fast response of the PMs make them ideal detectors for fast (subsecond) optical transients like pulsars.

In observations with ground-based γ -ray Cerenkov telescope in the range below ~ 30 GeV (where the cutoff of the so-called canonical pulsar is expected to be), the above discussed imaging method is not efficient anymore since the light yield is too faint to produce good images of extensive air showers. Without the γ /hadron discrimination capabilities provided by imaging, the background of cosmic-ray initiated showers becomes a serious problem, both because of its magnitude and because of the difficulty to estimate accurately its contribution to

the observed rate of events. But, in observing periodic sources, a timing analysis may in part overcome these difficulties. To perform a correct timing analysis, one needs to take into account the pulsar timing noise and glitches, which will be discussed in chapter 3. Contemporary radio pulsars ephemeris may be needed but one also can make use of the MAGIC central pixel

Chapter 2

Pulsars as sources of γ -rays emission

Eight pulsars have been detected at high energies so far, while most of the unidentified EGRET source are believed to be pulsars. One can distinguish two main types of pulsars, the **millisecond pulsars**, and the classical or **canonical pulsars**. The capability of MAGIC to detect either of these two types of pulsar will be investigated by ways of simulations. These simulations are used to derive the expected rates, detection times and the best pulsar candidates will be selected in function of detection capability.

2.1 Gamma-ray Pulsars: Canonical Pulsars & Millisecond Pulsars

The discovery of rotating neutron stars was achieved for the first time by a graduate student, Jocelyn Bell, on August 6, 1967 [49, Hewish et al., 1968]. She first noticed a peculiar train of radio signals when the sky at right ascension $19^h 19^{min}$ passed through the field of view. The transient periodic signal appeared again at about the same time of day. Furthermore a recording of the source with

sub-second time resolution on November 28, 1967, revealed pulses repeating at a regular period of 1.33 s. Pacini [80, Pacini, 1968] and Gold [43, Gold, 1968], one year later, arrived at the still valid interpretation: the pulsed signals were generated by rotating, magnetized neutron stars, and the radiation luminosity derives ultimately from rotational energy. Now, 35 years after the discovery of pulsars, about 1500 radio pulsars are known, and this number is increasing as more sensitive instruments come into operation. Neutron stars emit in several wavebands, as radio, infrared, X-ray and γ -rays (e.g. [37, Eikenberry et al., 1996], [85, Safi-Harb, 1996], [13, Carramiñana et al., 2000]).

The number of γ -ray pulsars has grown from two to eight over the five years since the launch of Compton Gamma Ray Observatory (CGRO) [95, Thompson et al., 1993], new data on these objects has resulted in many puzzles and provided some clues about their nature. But before going further in the different models for γ -ray pulsars and their EGRET characteristics, an overview of basic properties should be given.

2.1.1 Neutron Stars Basic Properties

Theoretical models of neutron stars [91, Shapiro & Teukolsky, 1983] show that the allowable range of masses is between ~ 0.2 and 3.0 solar masses ($M_{\odot}$), since a smaller mass would not provide enough gravitation to hold the star together in its condensed state, and a large mass would lead to further collapse into a black hole. They rotate with periods between a few milliseconds and a few seconds. The slower ones are called canonical pulsars ($P \geq 20$ ms), the faster rotators are called millisecond pulsars. Figure 2.1 displays the distribution of currently observed radio pulsars in a period vs. period-derivative ($P\dot{P}$) diagram [65, Manchester et al., 2001]. There are clearly two different groups of pulsars with respect to P and B_{ps} :

- (a) canonical pulsars for which $P \geq 20$ ms and $B_{ns} \geq 10^{10}$ G, for which the

spind down age $P/2\dot{P} \geq 3 \times 10^7$ years.

(b) millisecond pulsars for which the parameter conditions are reversed.

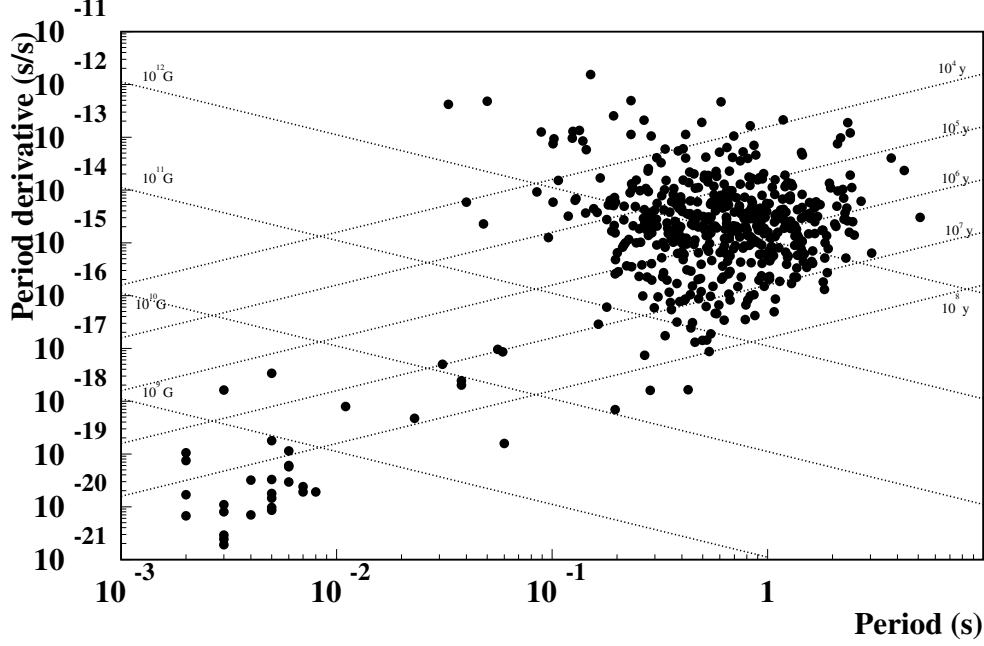


Figure 2.1: $P\dot{P}$ distribution of radio pulsar from the ATNF pulsar catalogue. The labeled lines indicates the "rotational" age ($\frac{P}{2\dot{P}}$) and the dipole field strength.

Conservation of angular momentum ($= I\omega$, where $I \propto Mr^2$ is the moment of inertia and ω the angular frequency, M is the mass and r is the radius of the neutron star) leads to the conservation of $r^2\omega$. A stellar core that collapses from a radius of $r_* \sim 10^{11}$ cm to $r_{NS} \sim 10^6$ cm will therefore increase its rate of rotation by $\sim 10^{10}$. A period of rotation of ~ 30 days will end up to periods of milliseconds. The interior of the star is a fully conducting medium (see Fig. 2.2). Hence magnetic flux (Br^2) will also be conserved during core collapse. This will increase the magnetic field from the typical value of a normal star to values of the order of 10^{12} G. The total rotational energy of a young neutron star ($\frac{1}{2}I\omega^2$) is of the order of 10^{51} ergs.

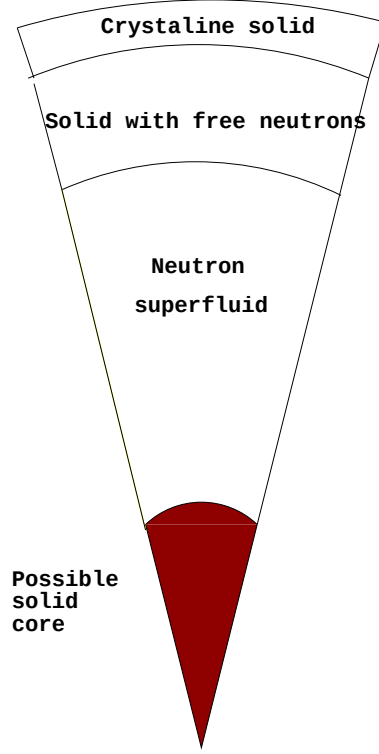


Figure 2.2: *Schematic view of a pulsar interior structure.*

Another estimate for the magnetic field strength can be obtained from the observed slowing down in the rotation. A spinning magnetic dipole radiates electromagnetic energy at a rate of $\frac{2}{3}\mu_{\perp}^2\omega^4c^{-3}$, where $\mu_{\perp} = \mu \cdot \sin\alpha$ is the perpendicular component of the magnetic dipole momentum $\mu \sim B_or^3$, while α is the angle between μ and the axis of rotation. Equating the dipole radiation losses and the slowdown rate of rotation, i.e.,

$$\frac{d}{dt}\left(\frac{1}{2}I\omega^2\right) = I\omega\dot{\omega} = -\frac{2}{3}\mu_{\perp}^2\omega^4c^{-3}, \quad (2.1)$$

with the perpendicular component of the magnetic dipole moment given by:

$$\mu_{\perp}^2 = -\frac{3Ic^3}{2}\frac{\dot{\omega}}{\omega^3} = \frac{3Ic^3}{2}P\dot{P} \quad (2.2)$$

Assuming a value of $I = 10^{45} \text{ g} \cdot \text{cm}^2$, and $\alpha \sim 90^\circ$, one obtains the following estimate for the polar field strength:

$$B_o \sim 3 \times 10^{19} \sqrt{P\dot{P}} \text{ Gauss} \quad (2.3)$$

Another important characteristic is the "rotational age" τ , which can be derived from the dipole radiation energy-loss rate under the assumption that $\mu_\perp$ remains constant (Eq. 2.2). It follows that $\dot{\omega} \propto -\omega^3$, which can be integrated to yield

$$\tau = \frac{\omega}{2\dot{\omega}} \times \left(1 - \left(\frac{\omega}{\omega_o}\right)^2\right) \quad (2.4)$$

where ω_o is the initial spin frequency. Since for most of the pulsars we expect $\omega_o \gg \omega$, the age reduces to $\tau \sim \omega/2\dot{\omega}$.

2.1.2 High Energy Emission: Polar Cap vs. Outer Gap Models

Pair creation paradigm is a pivotal element in any model of magnetospheric activity of rotation powered pulsars (RPP). The formation of electron-positron pairs ($e^\pm$ -pairs) are essential, since they are thought to be responsible for radio emission observed in radiopulsars, which is interpreted as the coherent curvature radiation from an $e^\pm$ plasma.

Pairs can be produced in magnetospheric environment via:

- photon absorption in a dense field of soft photons (photon-photon collision), provided that the center of mass energy is more then the electron rest mass.
- photon absorption in a strong magnetic field, provided that the threshold criterium has been met.

In either case a supply of high-energy (HE) photons is required in order to fulfill these stringent threshold conditions for pair creation. It is quite reasonable then

to assume that not all of those HE-photons would be subject to absorption. On the contrary, many HE photons will escape the magnetosphere without any attenuation. This argument leads us to expect that RPP (and all radiopulsars in particular) should be sources of HE radiation.

To make the production of HE photons possible, highly relativistic charged particles have to be injected into the magnetosphere. One may speculate that some of these particles will either retain their energy or regain it upon escaping from the source. It is up to theoretical models of the RPP activity to show whether the luminosity of HE radiation with associated beaming is favorable for detection with the IACT.

Although there are several explanations for pulsar γ -ray emission, two models are the most widely accepted among the pulsar community: the polar cap model and the outer gap model. As discussed above, neutron stars are natural unipolar inductors generating strong electric fields. These fields are able to pull charges out of the star against the force of gravity and it is believed that the charge density that builds up from the neutron star magnetosphere can short out the electric field parallel to the magnetic field ($E_{\parallel}$) everywhere except at a few locations. Such a non-zero $E_{\parallel}$ will accelerate charged particles, as discussed in Section 1.1.2. The location of these "gaps", where $\mathbf{E} \cdot \mathbf{B} \neq 0$, defines the main differences between the polar cap model (where the acceleration and radiation occur above the magnetic poles close to the neutron star surface) and the outer gap model (where the gap is located next to the cylinder of light, in the outer magnetosphere)[84, Rudenam & Sutherland, 1975], [15, Cheng & Ruderman, 1989].

2.1.3 Polar Cap Model

The magnetic field line structure of a neutron star can be described by a dipole. In polar coordinates its field lines satisfy the relation:

$$r \sin^{-2}\theta = R_o \quad (2.5)$$

where the dipole constant R_o defines a set of field lines which differ by azimuthal angle only. Polar caps are then defined as regions on the surface crossed by all lines for which the condition $R_o \geq R_{lc}$ is satisfied, where $R_{lc} = cP/2\pi$ is a light-cylinder radius for a period P . The field lines from the polar cap therefore close outside the light cylinder, whereas the field lines originating closer to the equatorial region close inside the light cylinder.

The polar cap radius will be given then by

$$r_{pc} = R_{ns} \times \left(\frac{R_{ns}}{R_{lc}}\right)^{1/2}, \quad (2.6)$$

with R_{ns} the radius of the neutron star. Beam particles (primary electrons) are ejected from the polar cap. They move along the open magnetic field lines. The accelerated electrons emit curvature γ radiation, due to the curved magnetic field lines, and non-resonant inverse Compton scattering on thermal photons, emitted by the hot surface of the neutron star.

These γ -rays emitted by curvature radiation eventually cross magnetic field lines and since the field is superstrong, magnetic pair production is unavoidable if the threshold criterium is met. These $e^\pm$ pairs emit γ -rays of higher energy due to inverse Compton scattering on thermal photons. The secondary soft $e^\pm$ resulting also radiate γ -rays (synchrotron radiation), but of lower energy compared to the primary energy. These secondary γ -rays pair-produce too, creating further pairs with even lower energy. This recursive process continues until the γ -ray energy drops below the threshold for pair creation, after which the electron-

photon cascade ceases, while the surviving γ -rays escape to infinity as an observed pulsed component [45, Harding & Muslimov], [36, Dyks et al, 2003].

The natural consequence of these polar cap processes is a superexponential cutoff E_o as discussed by Nel & de Jager [74, Nel et al., 1995].

Hence the spectrum of a pulsar (canonical or millisecond) is a superposition of these two effects and Compton upscattering (ICS). Bulik et al. [11, Bulik et al., 2003] show that a photon created with a momentum parallel to the local magnetic field line, at a height h above the neutron star surface, undergoes magnetic absorption if its energy satisfies approximately the following inequality:

$$E_o \geq 10^2 \left(\frac{P}{10^{-3}s} \right)^{1/2} \left(\frac{B_{ps}}{10^9 G} \right)^{-1} \left(\frac{R_{ns}}{10^6 cm} \right)^{-1/2} \left(1 + \frac{h}{R_{ns}} \right)^{5/2} GeV \quad (2.7)$$

where E_o gives the energy where the superexponential cutoff becomes important in the energy spectrum. P the periodicity of the pulsar and R_{ns} its radius.

2.1.4 Outer Gap Model

In competition with the polar cap model, is the outer gap model. It is based in outer gaps, whereby a potential drop $\Delta V \sim 0.2(\dot{E}/c)^{1/2} \sim B_s/P^2 > 10^{12} eV$ [15], with B_s the surface field strength and P the pulsar period, develops around the $\mathbf{E} \cdot \mathbf{B} = 0$ "null surface". The distance between this null surface, located near the light cylinder, and the pulsar is very large compared with the polar cap size. In this case we find that photon-photon pair production and the available acceleration potential determines the cutoff. Both the synchrotron and a VHE inverse Compton component can then escape from the outer gap to the observer, without conversion into pairs as a result of magnetic pair production. The reason for this is because the field strength near the light cylinder is diluted by the cubic law ($B_{lc} \sim B_o(\frac{R_{ns}}{R_{lc}})^3 \ll B_o$) relative to the stellar field strength B_o .

It is quite plausible that both the polar cap and the outer gap mechanism may

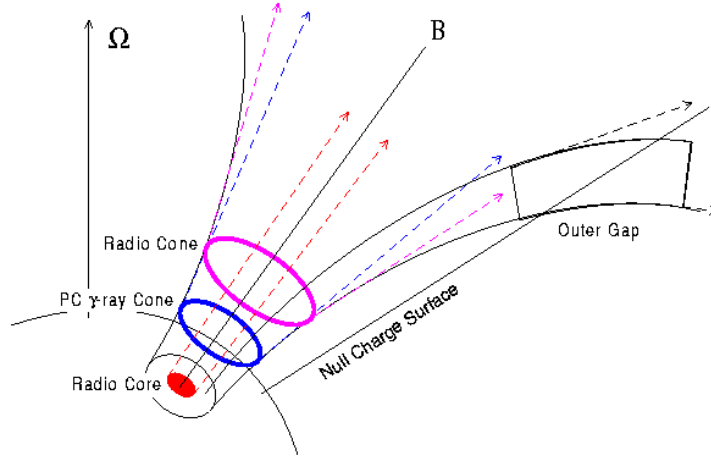


Figure 2.3: *Cartoon of polar cap, outer gap and radio beams. The outer gap is shown as a 2D projection ([46, Harding, 2002]).*

operate simultaneously (see Fig. 2.3), in which case a first order generic model (polar cap and/or outer gap) for the tails of pulsed differential spectra may be modeled as [25, de Jager, 2002] (see also eqn 2.9 for a more detailed discussion):

$$\frac{dN}{dE} = K_1 \left(\frac{E}{E_n} \right)^{\Gamma_1} \exp(-(E/E_{01})^b) + K_2 \left(\frac{E}{E_n} \right)^{\Gamma_2} \exp(-(E/E_{02})^c) \quad (2.8)$$

with E_{01} the cutoff corresponding to the polar cap process, whereas E_{02} would represent the characteristic outer gap cutoff energy. The second term would be absent in the case of pure polar cap γ -ray emitters, with the additional signature of a super exponential cutoff ($b \geq 1$ and $K_2 = 0$). An outer gap origin can be interpreted in terms of a non-zero K_2 , but with a slower roll-over ($b < 1$ and $c < 1$) compared to polar cap models, since the outer gap absorption process is controlled by photon-photon pair production, which is more weakly energy dependent than magnetic pair production above the polar cap [16, Cheng et al., 2000].

2.1.5 EGRET Characteristic of Pulsars

Seven or maybe eight γ -ray pulsars have been seen by the CGRO/EGRET instrument during its mission between 1991 and 1997, and several hard-spectrum unidentified EGRET sources have also been seen which are thought to be γ -ray pulsars for which the EGRET statistics are too weak to resolve any periodicity.

Fig.2.4 shows the light curves of γ -ray pulsars in several ranges of the electromagnetic spectrum. The correlation between radio and high-energy γ -ray profiles differs from pulsar to pulsar. For instance, the main γ -ray peak of the Crab is closely aligned with the radio peak, while Geminga is the unique example of a radio quiet pulsar. The observed γ -ray light curves show mostly bimodal emission patterns; only PSR B1706-44 has a more complex light curve. PSR B1509-58 was detected by the Burst and Transient Source Experiment (BATSE), the Oriented Scintillation Spectrometer Experiment (OSSE), and the Imaging Compton Telescope (COMPTEL) up to ~ 1 MeV. PSR B1706-44 and PSR B1055-52 [39, Fierro et al., 1993] show single broad pulse profiles above 300 MeV. The most recent pulsar discovery by EGRET is PSR B1951+32, which has a light curve similar to that of a Vela. So then, together with the Crab and Vela pulsar, a total of seven γ -ray pulsar are positively identified at the moment.

The Crab pulsar is the youngest and most luminous object amongst this sample. Its association with the Crab nebula is very clear, which serves as the "standard candle" for high-energy astrophysics. The coincidence in phase across the observed electromagnetic spectrum makes it also a good candidate for multiwavelengths calibration, where the optical central pixel (OCP) in MAGIC will be used to calibrate the γ -ray timing (see Chapter 5).

Although multi-wavelength spectra and light curves are clearly differences between these sources, some global trends can be derived from the consideration of spectral indices and the conversion efficiency of rotational energy into γ radiation.

The high energy spectra can be modeled [74, Nel & de Jager, 1995], using a

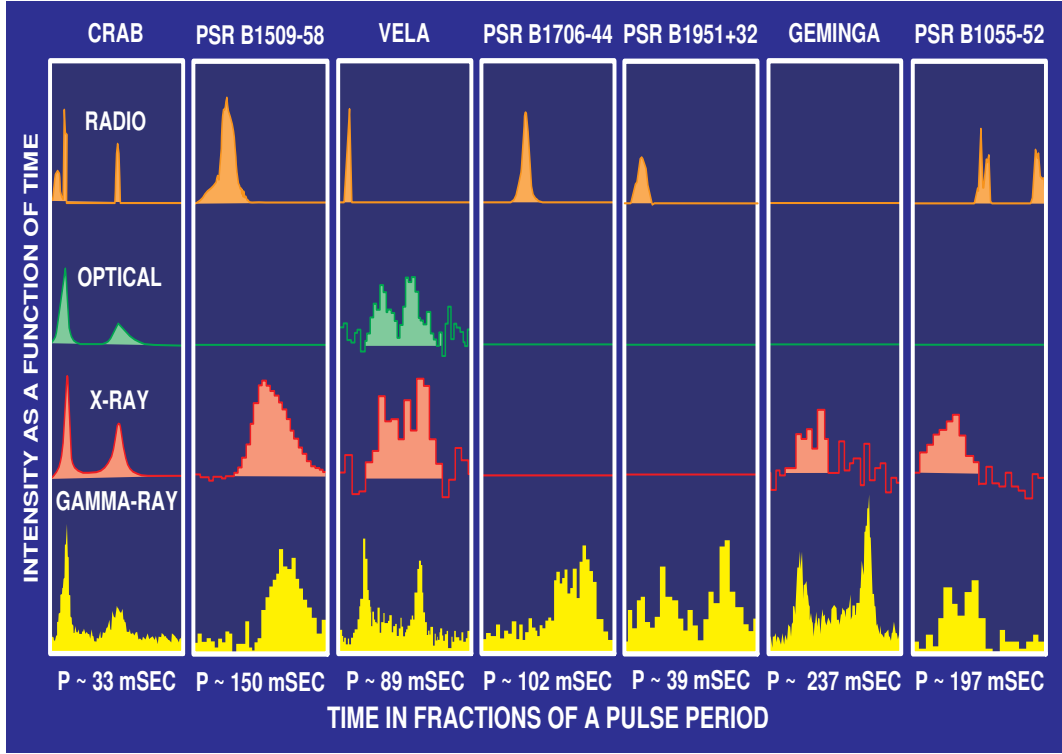


Figure 2.4: Pulsar detected by EGRET at high energy γ -rays, and their radio, optical and X-rays counterparts [75, Nolan et al., 1996].

conservative polar cap model, as

$$\frac{dN_\gamma}{dE} = K E^{-\Gamma} e^{-(E/E_o)^\alpha}, \quad (2.9)$$

where Γ is the photon spectral index in the highest energy range detected by CGRO (see Fig.2.5). The cutoff energy is given by E_o , whereas α determines the strength of the cutoff. The best fitted parameters are represented in Table 2.1.

A low cutoff energy makes detection with ground based telescope difficult, since the collection area $A(E)$ decreases with decreasing energy E , and a significant overlap of $A(E)$ and dN/dE would be required. We need then to decrease as much as possible the threshold of the telescope to be able to detect the pulsed emission. It requires a special pulsar trigger, at software level as well as at hard-

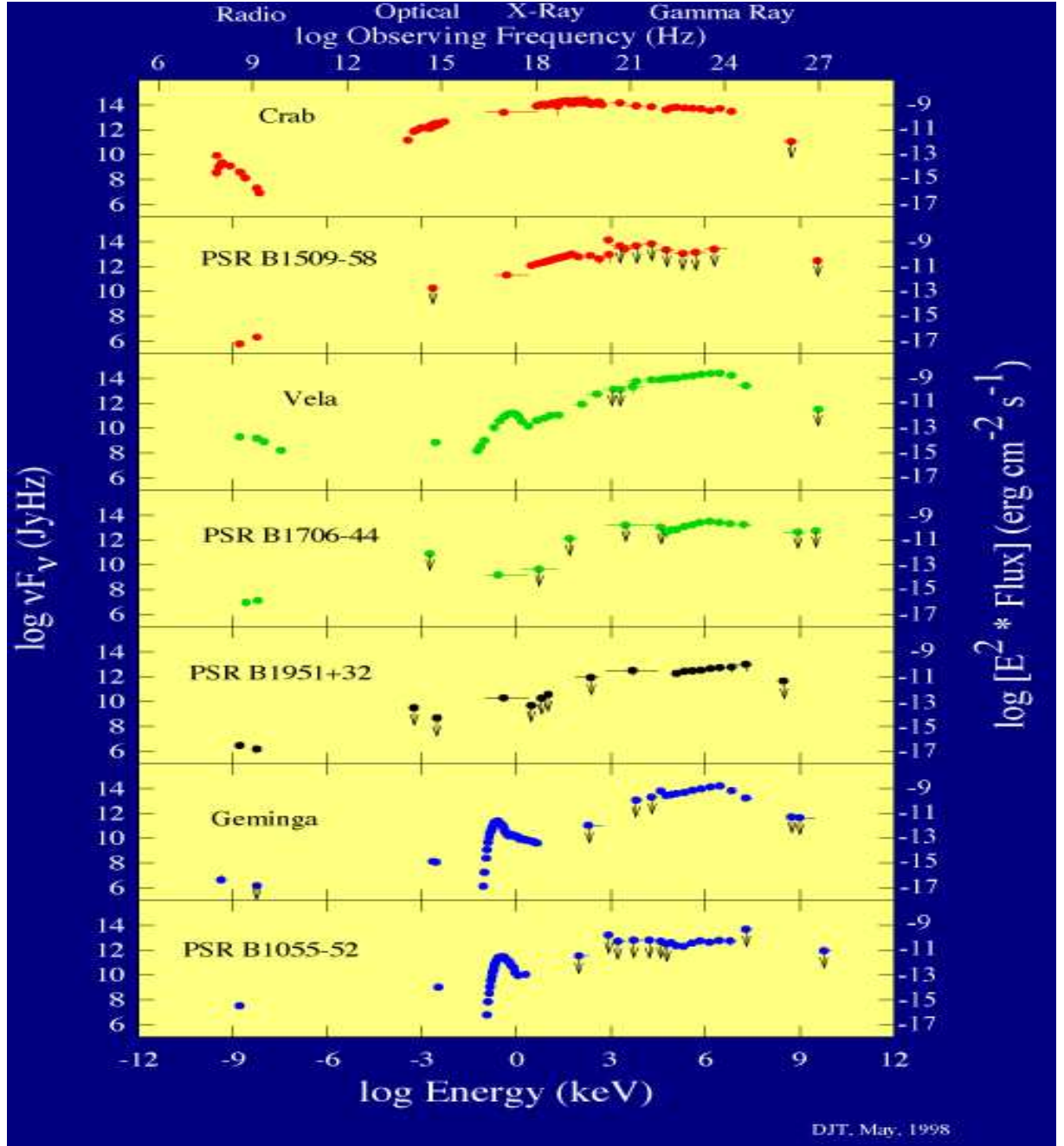


Figure 2.5: *Differential energy spectra (radio to γ -ray) of the seven CGRO pulsars. The fitted parameter using Eq. 2.9 are in represented in table 2.1 (Credit: D.J. Thompson (NASA/GSFC))*

Name	$k (\times 10^{-8})$ ($cm^{-2}s^{-1}GeV^{-1}$)	Γ	E_o (GeV)	α
Crab	24.0	2.08	30	2
Vela	138	1.62	8.0	1.7
Geminga	73.0	1.42	5.0	2.2
PSR B1951+32	3.80	1.74	40	2
PSR B1055-52	4.00	1.80	20	2
PSR B1706-44	20.5	2.10	40	2

Table 2.1: *Table of parameter fits for eq 2.9 [30, de Jager, 2002]*

ware level.

We will discuss the possibility of a canonical pulsar detection with the standard MAGIC operational mode. In the next chapters we will discuss a new technique to use the MAGIC telescope as a 5 to 10 GeV detector.

2.2 High-Energy Phenomena In Millisecond Pulsars

Beside the so called canonical pulsars, a sample of radio pulsars, the millisecond pulsars, have been detected with rotational periods in the range from 1 ms to 10 ms. These pulsars are rather old, and most probably have been recycled by some external mechanism (usually a binary system). The surface magnetic fields are lower than that of canonical pulsars, with $B_{ps} \sim 10^9$ G. Given such a low magnetic field, the cut-off energy of curvature photons due to pair production shifts to higher energies as we will discuss in the next section. The shift of the cutoff at higher energies, together with the relatively large collection area of MAGIC below 100 GeV, make these objects ideal candidates to be seen with MAGIC.

The majority of pulsars follow a course of evolution from birth in a supernova, through a slowdown from a period of some tens of milliseconds to a death at around one second, when the pair production rate becomes too low to sustain

radio emission. The millisecond pulsars constitute then a separate and much longer lived population. It is believed that they are the result of an interaction with binary partners. Most still having their companions, but some having lost them to become solitary millisecond pulsars.

The interaction between the binary partners, which leads to millisecond pulsars, is the transfer of mass. This is a process that can be observed from X-ray binary pulsars. Here the mass transfer provides the thermal energy for the intense X-rays; for pulsars, or millisecond pulsars, the essential parameter is the transfer of angular momentum from the binary orbit to the pulsar. The result is a spin-up to a rotation speed of some hundreds of rotations per second. Therefore they show up at the lower left corner in the $P-\dot{P}$ diagram.

The low magnetic field strength observed through Equation 2.3 also implies that some process of magnetic field decay must be responsible for the low magnetic field strength associated with millisecond pulsars.

Considering a most conservative polar cap, a sharp, super-exponential cutoff is still an inherent feature at the high energy end of the pulsar spectrum, so we use the following fit for the pulsed differential spectra:

$$E^2 \frac{dN}{dE} = K \cdot E^\alpha \cdot \exp\left(-\left(\frac{E}{E_o}\right)^2\right) \quad (2.10)$$

where E_o is the cutoff energy determined for magnetic absorption of the photons.

It was found [45, Harding et al., 2002] that in general the curvature radiation effect is very effective in screening the electric field above the pair formation front. When Inverse Compton Scattering screening does occur, it only screens the field locally.

According to the basic physics behavior of the pair production rate in the presence of crossed electric and magnetic fields ($\mathbf{E} \cdot \mathbf{B}=0$), one can calculate the

photon absorption rate [24, Daugherty & Lerche, 1975] and pair production. If we consider Sturrock's [94, Sturrock, 1971] model, pair production will occur whenever:

$$E_\gamma \cdot B_\perp \geq 10^{18.6}, \quad (2.11)$$

where $B_\perp$ is measured in Gauss and E_γ in eV. To see what is the relevance of the effect of curvature radiation in millisecond pulsars, one can express the maximum γ ray energy due to curvature radiation as a function of the surface magnetic field:

$$E_{\gamma \text{ max}} \simeq 10^{-77.5} B_{ps}^3 R_{ns}^{17/2} P^{-13/2} \quad (2.12)$$

where $E_{\gamma \text{ max}}$ is in eV, B_{ps} in Gauss, R_{ns} is the neutron star radius in cm and P the period in seconds. For canonical pulsars, with $B_{ps} = 10^{12} \text{ Gauss}$, $P=1 \text{ s}$, and $R = 10^6 \text{ cm}$, equation 2.12 yields

$$E_{\gamma \text{ max}} \simeq 3 \times 10^9 \text{ eV}, \quad (2.13)$$

And if we combine it with the condition of pair production then:

$$B_{\perp \text{ min}} \geq 10^9 \text{ Gauss} \quad (2.14)$$

For millisecond pulsars, due to the low surface magnetic field, the condition for the creation of pairs is not fulfilled, and the cutoff energy for curvature photon shifts to higher energies. Being conservative, we would take the slope of the spectrum α as 0 for $E > 10 \text{ GeV}$, compared to a more theoretical, but optimistic $\alpha = 2/3$, for unabsorbed curvature radiation in the absence of any pair production.

The γ -ray luminosity derived from the rotational energy loss rate ($\dot{E}$) for millisecond pulsars is also affected by the special characteristic of these type of objects.

A good approximation is to assume that $L_\gamma \approx 0.5L_{prim\ max}$ where $L_{prim\ max}$ is the maximum power of the primary electron beam in the polar cap pulsar model. Following the model given by Harding et al., [45, Harding et al., 2002] the luminosity of the primary beam is:

$$L_{prim} = \alpha_r c \int |\rho_e| \Phi dS \quad (2.15)$$

where ρ_e is the value of the electron charge density, $\Phi(z, \xi, \phi)$ is the electric potential as a function of the altitude above the polar cap (h), the magnetic colatitude of a field line scaled by the magnetic colatitude of the last open field line is named (ξ), and the magnetic azimuthal angle is ϕ . Here $\alpha_r = (1 - r_g/R_{ns})^{1/2}$, r_g is the gravitational radius of a neutron star. The integral must be calculated over a spherical surface dS cut by the last open field lines at the radial distance $r = R_{ns}\eta$.

Assuming that no pairs are produced or at least the electric field screening is very low, we can find an analytical expression for the electric potential Φ , which corresponds to the saturated and unscreened regime of acceleration of primaries, and performed integration over ξ and ϕ in Eq. 2.15 in function of $\eta = r/R$, where r is the radial distance and χ the pulsar obliquity:

$$L_{prim}(\eta)|_{\cos\chi \approx 1} = \frac{3}{4} \kappa (1 - \kappa) \left(1 - \frac{1}{\eta^3}\right) \dot{E}_{rot} \quad (2.16)$$

where κ represents the dependence on neutron star radius and momentum of

inertia I:

$$\kappa = 0.15 \frac{I}{10^{45}} \left(\frac{R_{ns}}{10^6} \right)^{-1} \quad (2.17)$$

Then for a maximum radial distance ($\eta \rightarrow \infty$)

$$L_{prim,max} = \frac{3}{4} \kappa (1 - \kappa) \dot{E}_{rot} \quad (2.18)$$

Therefore, defining $I_{45} = I/10^{45}$ and $R_6 = R_{ns}/10^6$ we can write the millisecond and γ -ray luminosity as:

$$L_\gamma \approx 0.05 \frac{I_{45}}{R_6^3} \left(1 - 0.15 \frac{I_{45}}{R_6^3} \right) \dot{E}_{rot} \quad (2.19)$$

This means that the maximum efficiency of conversion of pulsar spin-down power into the high-energy quanta may amount to 10%.

The γ -ray luminosity expected from millisecond pulsars is very low compared with canonical pulsars, but according to equation 2.7 they are expected to radiate mostly in the range between 10 GeV and 1.5 TeV, due to the unscreening effect which shifts the cutoff to higher energies. Moreover, the low γ -ray luminosity is offset by the large collection area of the next generation of telescopes such as MAGIC which makes the detection of these objects feasible.

2.3 The Detection of Pulsars with MAGIC

To investigate the expected rate of pulsed events in the MAGIC telescope it is necessary to know not only the spectrum of the source but also the effective area of the telescope:

$$R_\gamma = \int_{E_{th}} \frac{dN}{dE} A_{eff}(E) dE \quad (2.20)$$

The spectrum $\frac{dN}{dE}$ is modalized using Eq. 2.9 as calibrated with table 2.1 for canonical pulsars, while for millisecond pulsars Eq. 2.19 gives us the expected spectrum.

On the other hand, to calculate the effective area, one has to simulate electromagnetic and hadronic cascades, by Monte Carlo Simulations, as well as the response of the telescope (reflector and camera).

2.3.1 Simulation Programs

There are several programs which perform a simulation of the gamma/cosmic ray development in the atmosphere (KASKADE, MOCCA...). The MAGIC collaboration uses CORSIKA (COsmic Rays SIMulation for KAscade [53, Knapp et al., 1998]).

CORSIKA is a detailed Monte Carlo program that tracks the evolution of EAS in the atmosphere initiated by photons, protons, nuclei, or any other particle. The program simulates the transport of particles through the atmosphere and their decays and interactions with air targets. It includes also a model of the atmosphere and takes into account the Earth's magnetic field. All particles are tracked along their trajectories and the output of the program gives type, energy, location, direction and arrival times of all secondary particles that are created in an air shower and pass a selected observation level.

Once the shower simulation is complete, the response of the telescope is simulated using two programs, the "reflector" and the "camera" programs, both developed by the MAGIC collaboration.

With the reflector program (authors J.C. Gonzalez, D. Bastieri & C. Bigongiari

and A. Moralejo [72, Moralejo, 2002]) it is possible to track the Cerenkov photons generated by CORSIKA up to the camera plane. The camera program (authors J.C. Gonzalez, O. Blanch, D. Petry & H. Kornmayer) calculates the response of each photo multiplier of the camera to the Cerenkov photons, propagates the signal through the camera and optical fibers and simulates its digitization FlashADCs. The result is a number of ADC counts per time slice in each FlashADC channel. The program also simulates the two levels of trigger.

2.3.2 MAGIC Effective Areas

For γ -ray showers, the simulation is performed in the 10 GeV to 30 TeV energy range, which corresponds with a background of proton showers with energies between 30 GeV to 30 TeV. Given the fact that protons produce several times less Cerenkov photons than γ -rays of the same energy, we have to simulate protons with corresponding higher energies, even more since a large number of proton showers developing far away from the pointing direction with high energy also contribute to the signal collected. A special set of γ -ray data for pulsars studies has been run in Madrid, under the supervision of R. de los Reyes, with energies between 5 GeV to 10 GeV. These types of shower deserve a different treatment which will be discussed in subsequent chapters.

In the camera simulation, the trigger condition comprising any 4 neighboring channels (4NN) in a close-packed configuration above the pre-set threshold of 3.5 ph.e.s together, has been applied both for γ -rays and protons (in a close-packed configuration each fired pixel has at least two fired neighboring pixels).

Simulated night sky background, (or Light of Night Sky, LONS) has been included as well, the intensity of which has been measured at La Palma, at the site of the HEGRA experiment. The LONS intensity I_{LONS} [68, Mirzoyan et al., 1994] for a narrow angle detector in the 300-600 nm range has been measured to

be:

$$I_{LONS} = (1.7 \pm 0.4) \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \text{s}^{-1} \text{sr}^{-1} \quad (2.21)$$

This value is a mean intensity, valid for extragalactic sources. A new value was calculated using the central pixel of one of the HEGRA telescope, CT1, as discussed in Chapter 4.

In our simulation we assumed a point source of γ -rays with flux:

$$\frac{dN}{dE} = 4.7 \times 10^{-7} \left(\frac{E}{1 \text{ GeV}} \right)^{-1.5} \text{ ph} \cdot \text{cm}^{-2} \text{s}^{-1} \quad (2.22)$$

and for protons, we used the known proton spectrum:

$$\frac{dF_p}{dE} = 1.95 \times 10^4 \left(\frac{E}{1 \text{ GeV}} \right)^{-2.75} \text{ m}^{-2} \text{s}^{-1} \text{sr}^{-1} \text{GeV}^{-1} \quad (2.23)$$

Although this contribution is not as important as the proton background, helium nuclei and electrons [102, Wiebel-Sooth, 1998] were also simulated to include them as a background. The spectra have been measured to be:

$$\frac{dF_{He}}{dE} = 5.22 \times 10^3 \left(\frac{E}{1 \text{ GeV}} \right)^{-2.63} \text{ m}^{-2} \text{s}^{-1} \text{sr}^{-1} \text{GeV}^{-1} \quad (2.24)$$

$$\frac{dF_{e^-}}{dE} = 5.72 \times 10^2 \left(\frac{E}{1 \text{ GeV}} \right)^{-3.26} \text{ m}^{-2} \text{s}^{-1} \text{sr}^{-1} \text{GeV}^{-1} \quad (2.25)$$

We calculate then the trigger efficiency for a given energy and impact parameter defined by:

$$\epsilon(E, r) = \frac{\text{events accepted by trigger}}{\text{total number of events}} \quad (2.26)$$

The simulation was repeated for different zenith angles. For γ -ray showers the effective area is given by the integration of the trigger probability at an impact parameter defined by:

$$S_\gamma(E) = \pi \sum_{r_i=0}^{r_{max}} \epsilon(E, r_i) [(r_i + r_o)^2 - (r_i - r_o)^2] \quad (2.27)$$

Figure 2.6 shows the γ -ray effective area for different zenith angles [32, de los Reyes et al., 2002], which was parameterized for further studies using an analytical function.

For the calculation of the effective area for protons, we have to take into account that propagation through the interstellar magnetic field makes the proton background isotropic. Therefore the effective area has been calculated as:

$$S_{had} = \int d\varphi \int r \cdot dr \int d\phi \int \sin \delta d\delta \cdot \epsilon(r, \delta), \quad (2.28)$$

where the coordinates with respect to the telescope axis are explained in Fig. 2.7.

Fig. 2.8 [32, de los Reyes et al., 2002] shows the effective area for protons for several incident zenith angles.

Both areas tend asymptotically to $\sim 10^5$, although the slope is different. If additional cuts are imposed to improve the γ /hadron separation, Fig. 2.8 will be modified.

The protons rate, for no γ -hadron separation, based on 4NN trigger, using the background flux from Eq. 2.23 and their effective area (Fig.2.8), is given by

$$R_h = \int_{30 \text{ GeV}} 1.94 \times 10^4 \cdot E^{-2.75} S_{had}, \quad (2.29)$$

Integrating for all angles, from 30 GeV to 30 TeV, we obtain a total rate of protons of $\sim 230 \text{ Hz}$.

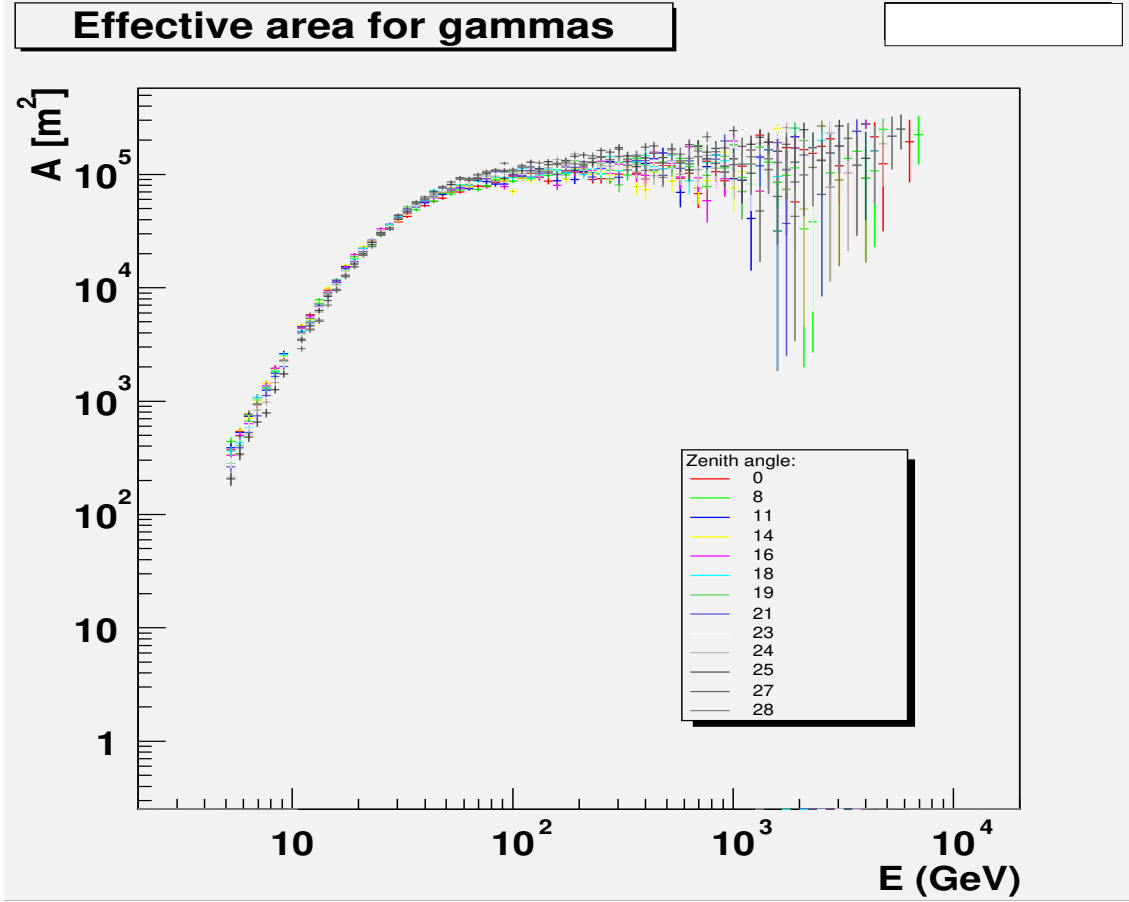


Figure 2.6: *Effective collection area after triggering for γ -ray for a camera with a 2.2° trigger diameter. The trigger condition "any 4 neighboring pixels above the threshold of 4 ph.e." has been applied.*

With the known rate of protons, using the EGRET spectrum 2.1 for canonical pulsars and Eq. 2.19 we can estimate the expected rates of the γ -ray sources for a given threshold.

We should define first the basic scaling parameter ([26, de Jager et al., 2001]), for all statistics, which measure the integrated difference between the measured pulse profile and a uniform distribution:

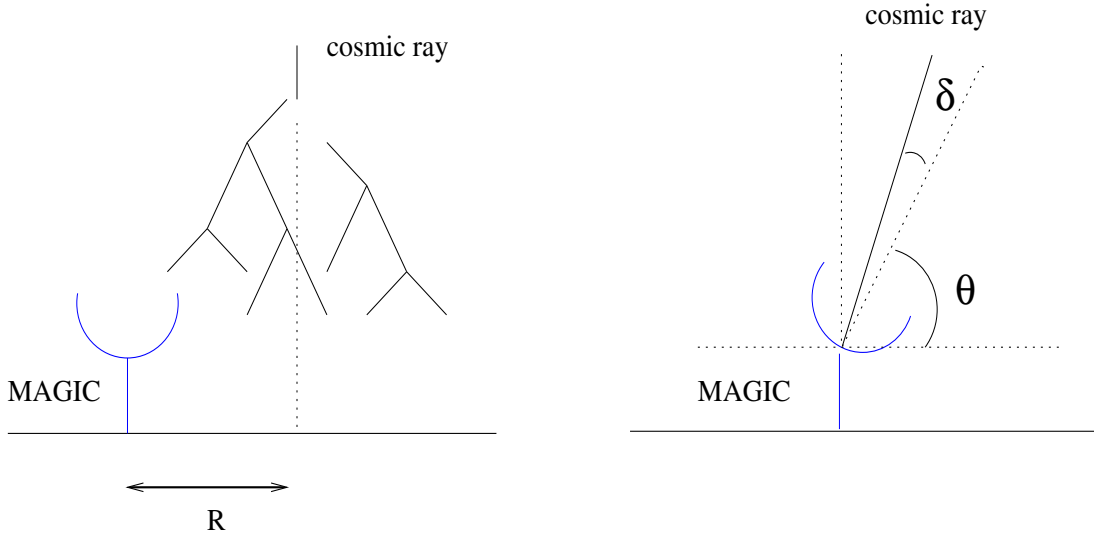


Figure 2.7: Scheme of the coordinates used to calculate the effective area for an isotropic background. The plot on the left illustrate showers far away from the telescope but that still produce light on the camera, being R the impact parameter. The figure on the right shows the zenith and azimuth angle with respect to the arrival direction.

$$x = Q \cdot p\sqrt{N} \quad (2.30)$$

Where $p = R_{pulsed}/(R_p + R_{pulsed})$ is the pulsed fraction and the total number of detected events is given by $N = (R_{pulsed} + R_p) \cdot T$, with T the observation time, R_{pulsed} the pulsed signal rate and R_p the background rate.

The quality factor Q describes the efficiency of the cut against background events and acts to improve the detection significance. The quality factor Q is assumed to be unity in this discussion, but can be improved to $Q \geq 1$ if we assume that the γ -ray spectra cut off above the energies E_o as listed in table 2.1 for typical canonical (EGRET) pulsars. In this case one can impose the restriction of $E_\gamma \leq E_o$, which implies a SIZE cut against showers with sizes larger than E_o . This will automatically reject proton showers with sizes larger than the selected

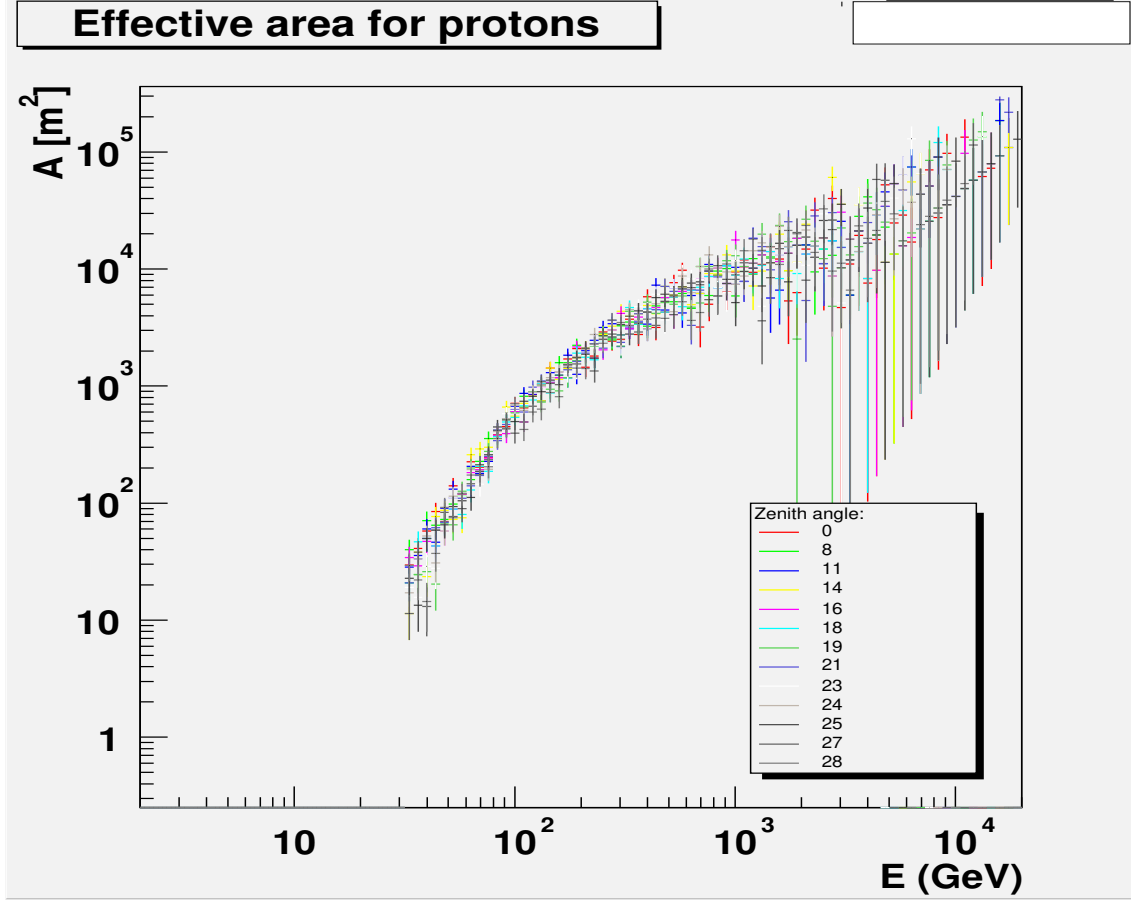


Figure 2.8: *Effective collection area after triggering for protons for a camera with a 2.2° trigger radius. The trigger condition "any 4 neighboring pixels above the threshold of 3.5ph.e." has been applied, and represents the full collection area before any cuts have been applied.*

threshold, in which case the collection area S_p , as shown in Figure 2.8, will be overall reduced such that the proton rate R_p will also be reduced. This, in effect, increases Q in Eq. 2.30. A further improvement can be achieved if the shape and concentration of such low energy showers are further exploited to increase Q . The latter was investigated, but only moderated Q -factors of no more than 3 were achieved. Rather than exhausting this venue, a totally new approach to achieve relatively large collection areas below 10 GeV will be explored in Chapter

2.4 MAGIC Detection Sensitivity for EGRET & Millisecond Pulsar63

5.

For the remainder of this chapter, a most conservative factor $Q=1$ will be assumed to show what the minimal detection power of MAGIC will be to detect EGRET and millisecond pulsars.

2.4 MAGIC Detection Sensitivity for EGRET & Millisecond Pulsars

If no γ /hadron discrimination is applied, that is, the so-called quality factor Q is equal to 1:

$$T_{obs} = \left(\frac{\sigma}{Q}\right)^2 \frac{R_h}{R_\gamma^2} \quad (2.31)$$

The required observation time depends on the threshold of the telescope as represented in table 2.2. In table 2.2 we present the observation times assuming a energy threshold of 5 for 6 pulsars, for an ideal effective area without added noise that modified the curve given by 2.6. We also calculate the rates and required times for a given threshold of 30 GeV , although we are aware that the effective area curve will change. A significance in the signal of $\sigma = 3$ is required.

Pulsar	Rate(Hz) for 5 GeV	Rate(Hz) for 30 GeV	T_{obs} (hours) for 5 GeV	T_{obs} (hours) for 30 GeV
Crab	1.45	0.25	0.27	9.0
Vela	0.45	2×10^{-4}	2.81	2×10^7
Geminga	0.24	9×10^{-24}	10.18	— — —
PSR B1951+32	0.90	0.31	0.70	5.9
PSR B1055-52	0.09	0.012	72.40	4×10^3
PSR B1706-44	0.74	0.39	1.03	3.8

Table 2.2: *Expected rates and observation times for EGRET pulsars at for a MAGIC threshold of 5 GeV and 30 GeV*

The cutoff energy restricts the canonical pulsar detection since the γ -ray flux

drop exponentially from E_o . Therefore, pulsars like Geminga, with a cutoff energy at $E_o = 5 \text{ GeV}$, are very difficult to detect, while Crab will be the best candidate. Nevertheless, applying the technique explained in Chapter 6, we will have difficulties in the Crab pulsar detection while Geminga will become the best candidate.

For the millisecond pulsars, a catalogue of radio pulsars [65, Manchester et al., 2001] with rotational periods faster than 30 ms has been investigated to establish if MAGIC will be able to detect their γ -ray emission at energies between 30 GeV and 1 TeV, within the context of polar cap emission from rotationally powered pulsars. The expected γ -ray rate and the time required to detect a signal with 3σ significance have been calculated for a total of 73 millisecond pulsars.

We have calculated the spectrum of these 73 millisecond pulsars using expressions discussed in the previous section, including the cutoff energy given by 2.7, based on observational parameters. To calculate the constant k we take into account the relation that predicts that the high energy luminosity L_γ is, in a first approximation, $\sim 5\%$ of the pulsar spindown power $\dot{E} = 4\pi^2 I \dot{P} / P^3$, which can be derived from basic pulsars parameters obtained with radio observations.

$$L_\gamma = \int_0^\infty E \frac{dN}{dE} dE = 0.05 \dot{E} \quad (2.32)$$

Assuming that a general value of $B_{ps} = 10^9 \text{ G}$, $R_{ns} = 10^6 \text{ cm}$ and $R_{ns} \gg h$. Out of the spectra and the effective area of MAGIC, we can derive the expected observation time for a detection (assuming no further techniques to clean the image and reject the background).

For each pulsar the expected integral rate of high energy photons (between 10 GeV and 1 TeV) is calculated. Using again the rate of hadrons ($\sim 400 \text{ Hz}$) and the MAGIC effective area for γ -rays we can derive the time required to have a detection of 3σ significance. Table 2.3 shows the millisecond pulsars which can

2.4 MAGIC Detection Sensitivity for EGRET & Millisecond Pulsar65

Table 2.3: List of ms pulsars detectable with MAGIC within less than 48 hours.

Pulsar name	P(s)	$\log(B_{sp})$ (G)	$\log(\dot{E})$ ($erg \cdot s^{-1}$)	E_o (GeV)	dist (kpc)	z($^{\circ}$)	$T_{obs}^{\alpha=0}$	$T_{obs}^{\alpha=2/3}$
J0437-4715	0.00576	8.76	34.08	418	0.142	75	270 s	0.036 s
J1744-1133	0.00407	8.28	33.72	1058	2.0	39	0.89 h	0.20 s
J2124-3358	0.00493	8.50	33.83	698	0.248	61	1.9 h	0.60 s
J1909-3744	0.00295	8.32	34.35	826	0.553	65	4.1 h	1.1 s
J1807-2459	0.00306	9.16	35.98	119	3.27	52	6.1 h	8.9 s
J0030+0451	0.00487	8.34	33.54	1000	0.235	24	12 h	2.7 s
J1824-2452	0.00305	9.35	36.35	79	5.5	52	14 h	31 s
J1959+2048	0.00161	8.21	35.20	770	1.53	8	15 h	5.3 s
J1024-0719	0.00516	8.49	33.73	730	0.354	35	20 h	6.2 s
J1300+1240	0.00622	8.93	34.27	296	0.624	16	33 h	24 s
J0034-0534	0.00188	7.99	34.48	1405	0.98	33	37 h	6.5 s

be detected with MAGIC within 48 hours as discussed previously. Table 2.3 lists the value of the period P, magnetic field at the pulsar surface B_{ps} , spindown power $\dot{E}$, cutoff energy E_o using equation (2.7), distance d of the pulsar in Kpc, zenith angle at culmination at the MAGIC site in La Palma $|l - \delta|$ (where l is the geographical latitude at La Palma and δ is the pulsar declination) and the detection time for millisecond pulsars, assuming $\alpha=0$ and $\alpha=2/3$. A total of 11 pulsars are found. Since in the most pessimistic scenario a minimal value of $Q \sim 3$ is expected from MC, this number would increase to a few more millisecond pulsars but the detection time would be reduced to ~ 1 h for most of them.

Whereas MAGIC presents the first opportunity by a ground based telescope to detect the fast spectral turnovers from canonical pulsars, the relatively large energy turnovers expected from millisecond pulsars imply that previous experiments operating at higher energies may have had the sensitivity to detect these fast rotators.

Table 2.4: *Upper Limits to the flux set up by the Whipple telescope. The flux upper limit is given in units of $10^{-11} \text{cm}^{-2} \text{s}^{-1}$*

Pulsar name F(E>500 GeV)	Type	Flux Upper Limit
J1012+5307	MRB	<1.13
J1300+1240	MRB	<1.13
J1537+1155	MRB	<1.14
J1959+2048	MRB	<1.55
B1855+09	MRB	<1.4

2.4.1 Millisecond Pulsar Observation by Previous Second Generation γ -ray Telescopes

Several millisecond pulsars have already been observed by different γ -ray observatories. Some of these millisecond pulsars were claimed to be γ -ray emitters detected by the Durham group with the telescope Mkn 6 [9, Bowden C.C.G. et al., 1991]. Nevertheless no further confirmation was done with more sensitive telescopes such as Whipple. The 10 m optical reflector of the Whipple collaboration observed millisecond pulsars during two-three years until 1999 [44, Hall et al., 1999]. No significant excesses were detected either in unpulsed, steady emission or pulsed emission in any of these objects during several years, but upper limits to the flux were derived for a energy threshold larger than 500 GeV. Table 2.4 shows the flux upper limits for the pulsars observed by Whipple.

According to our results based on a conservative polar cap model (table 2.3), two of the ms pulsars observed by Whipple have an expected cutoff energy lower than the energy threshold of the detector (500 GeV), making the detection improbable. Similarly, PSR J1300+1240 has a cut-off energy at around 300 GeV, whereas a cutoff of ~ 60 GeV is expected for the millisecond radio binary pulsar PSR J1537+1155. The non-detection of pulsed γ -rays from these millisecond pulsars is therefore not in contradiction with the polar cap model. The flux up-

2.4 MAGIC Detection Sensitivity for EGRET & Millisecond Pulsar 67

per limits from the other three pulsars will be used to see if the non detection is consistent with the luminosity predicted by Eq. 2.19.

$$F_{up} \geq \int_{E_{th}}^{\infty} K \cdot E^{-2} \exp\left(-\left(\frac{E}{E_o}\right)^2\right) dE \approx \int_{E_{th}}^{E_o} K \cdot E^{-2} dE \quad (2.33)$$

The spectral index of E^{-2} follows from expected modifications as discussed by Dyks et al. [36, Dyks et al., 2003]

Hence, the value of the constant K

$$K \leq F_{up} \left(\frac{E_{th} E_o}{E_o - E_{th}} \right) \quad (2.34)$$

The pulsar spectra dN/dE can be now calculated with the new value of K derived from the flux upper limit, and then the expected luminosity (in ergs/s) of millisecond pulsars is given by:

$$L_{\gamma} \leq \Delta\Omega \cdot d^2 \int E \frac{dN}{dE} dE = \omega \dot{E} \quad (2.35)$$

Where the factor $\Delta\Omega \geq 1 - 3$ sr is the solid angle of γ -ray emission, and d is the distance to the pulsar in cm.

If the value of ω , given by the flux upper limit, is above the limit $L = 0.05\dot{E}$, then a telescope with a threshold energy of 500 GeV (e.g. Whipple) would not be able to detect the pulsed emission, while if $\omega \leq 0.05$ then Whipple was sensitive enough to have detected it and it would mean that the theory in Section 2.2 overpredicts γ -ray emission from pulsars.

Harding et al. [45, Harding et al., 2002] did the same calculations for theoretical values of several pulsars. We have tested our method with the three ms pulsars observed by Whipple, and found that all them have values of $\omega \geq 0.05$. This is

then consistent with the polar cap theory, for which predictions for MAGIC was presented.

Fig. 2.9 shows the relation between the spin down power $\dot{E}$ and Luminosity as derived from General Relativistic constraints (GR) and the Whipple upper limits interpreted by the GR model.

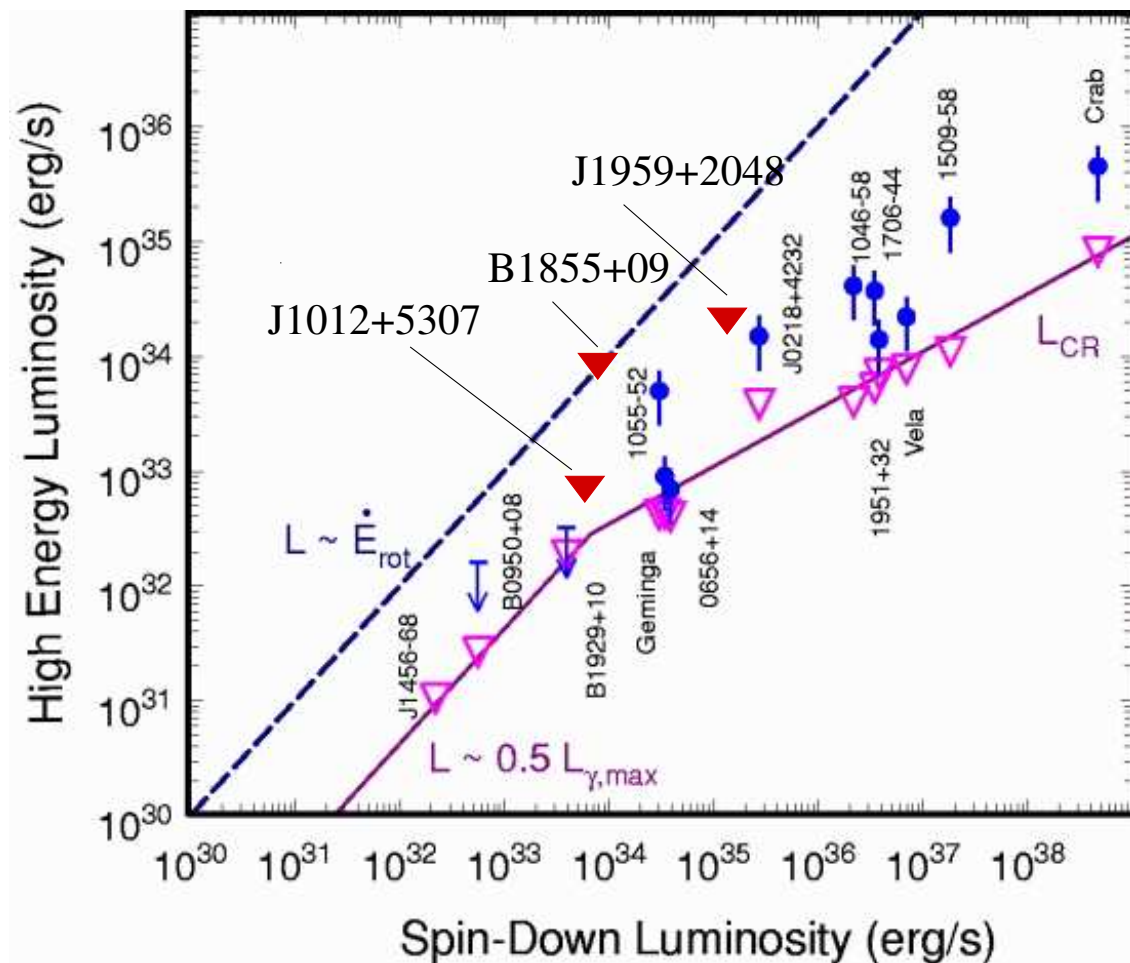


Figure 2.9: *Predicted and observed high-energy luminosity vs. spin down energy. The filled circles are the luminosities of the detected γ -ray pulsars, the open triangles are predicted luminosities [45, Harding et al., 2002], and the green triangles are our predictions for the Whipple ms pulsars.*

The solid curve is the theoretical prediction. Below $\dot{E} \leq 10^{34} \text{erg} \cdot \text{s}^{-1}$ the luminosity for the unscreened case is apply ($L_\gamma \sim 0.05\dot{E}$), and from here the

curvature radiation is taken into account when calculating the luminosity of the primary electron beam.

It is clear that the Whipple predictions are well above the theoretical prediction, illustrated by the solid line in Fig. 2.9.

2.5 Discussion

We explored the potential of the MAGIC telescope to detect GeV to multi-GeV pulsed emission from canonical and millisecond pulsars. To date no confirmed pulsed γ -ray emission from radio pulsars has been seen from any ground-based γ -ray telescope. The large mirror collection area allows MAGIC to go further down in threshold energy, compared to other experiments such as H.E.S.S. [51, Hoffman et al., 2001] or VERITAS [99, Wakely et al., 2003] arrays. Based in its collection efficiencies and the EGRET characteristics of some canonical pulsars, we showed that MAGIC can detect pulsations from EGRET pulsar at energies around ~ 30 GeV.

We show some prevision of required observation time. The Crab pulsar is the best candidate because its location with respect to La Palma, its high cutoff energy and intense flux, and it can be detected within a night. PSR B1951+32 is a good candidate that can be detected within 7.5 h with 3σ significance.

Due to the lower magnetic field in the polar cap of ms pulsars, $B_{sp} \sim 10^8$ to 10^9 G in contrast to the range of $B_{sp} \sim 10^{11}$ to 10^{13} G for canonical pulsars, the ms pulsar cutoff energies are shifted to higher values, up to ~ 1 TeV. Such a high energy cutoff favors the detection with MAGIC, even in the case that the γ ray flux is low, since the telescope reaches its large collection area $\sim 10^4$ m² at energies above 30 GeV.

Assuming a flux of γ -rays from only curvature radiation in absence of any pair production, that is, a spectral index of $\alpha=2/3$, the observation times $T_{obs}^{\alpha=2/3}$ become much smaller than the required times when electron/positron pairs do

not screen the radiation. In a more realistic point of view, using a spectral index $\alpha=0$, the required observation time increases. From the times given in 2.3, we can select the best ms pulsars to be observable by MAGIC.

To select the best targets, we also have to take into account the zenith angle range of the source at the MAGIC site. Therefore PSR J1959+2048 and PSR J0030+0451, with cutoff energies of ~ 770 GeV and ~ 1 TeV, would be our best choice for MAGIC during phase I. Predictions show that pulsar PSR J1959+2048, with a γ -ray luminosity of

$$L_{\gamma} = 1.027 \times 10^{34} \text{ erg/s} \quad (2.36)$$

should be detectable in ~ 15 h, and its zenith angle at upper culmination is 8° , while PSR J0030+0451 would be seen in ~ 1 day. On the other hand, the cutoff is only at a hundred GeV, which that means effective proton rejection at these energies can be efficiently performed using the conventional Hillas parameter approach. If we assume a quality factor of $Q=3$ the detection times reduces to a few hours in the first case and a few minutes in the second.

High energy pulsed emission from millisecond pulsars has not been detected by any Cerenkov or γ -ray detector, although pulsar theory predicts it. Therefore it is clear that the detection of pulsed emission from millisecond pulsars up to a few hundred GeV, or at least to the values predicted in table 2.3 should shed new light on the acceleration and γ -ray production process in millisecond pulsars.

Chapter 3

The Optical Central Pixel

Imaging Atmospheric Cerenkov Telescopes (IACT) can be used not only to detect γ -rays, but also to detect the optical emission of an astronomical object through the increase it generates in the currents of the camera pixels. Each of these pixels has a typical field-of-view on the order of $0.1\text{-}0.3^\circ$. Hence, when an IACT follows a typical astronomical object, its optical emission is fully contained inside the telescope central pixel.

The poor angular resolution of imaging γ -ray telescopes is offset by the large collection area of the next generation telescopes such as MAGIC (17 m diameter), which makes the study of optical emission associated with γ -ray sources feasible. In particular the optical photon flux causes an increase in the photomultiplier DC currents. Furthermore the extremely fast response of photomultipliers (PMs) makes them ideal detectors for fast (subsecond) optical transients and periodic sources like pulsars.

The optical pulsed emission of the Crab pulsar has already been detected using IACTs [89, Srinivisan et al., 1997] and other Cerenkov detectors [34, de Naurois et al., 2001]. Since the first optical observation of the Crab pulsar in the late 1960s [18, Cocke et al., 1969] only 4 more isolated pulsars have been seen to pulsate optically: PSR B0833-45 in Vela [98, Wallace et al., 1977], PSR 0540-

69 in the LMC [70, Middleditch et al., 1985] and more recently PSR B0656+14 [87, Shearer et al., 1997] and Geminga [88, Shearer et al., 1998]. The main purpose of such studies with Cerenkov telescopes are usually to determine the timing parameters of the pulsar with which the γ -ray arrival times can be folded, thus reducing the search for pulsed emission to a single statistical trial. It also allows us to test the accuracy of the timing system and solar system barycentric correction software when comparing the timing parameters against, for example, contemporary radio data.

This Chapter will, however, exploit the timing signature of the pulsar, using the central pixel of CT1, as well as DC observations of the bright nearby standard star ζ Tauri, to derive from the light of the night sky (LONS) around the Crab Nebula. The difference between these two numbers should give us an indication about the consistency in the LONS as derived using two different techniques: temporal versus spatial, convolved with the differences expected on a scale size of a few pixels due to galactic structure.

The observed rate for CT1 pulsations can also be used to obtain the expected response for the much larger MAGIC telescope. Furthermore, since the astronomical community does not monitor the Crab pulsar continuously in radio or optical, such supportive observations with Cerenkov telescopes may also help to serendipitously discover pulsar glitches or period irregularities in the future, given the relatively long integration times spent to obtain calibration runs for the spectrum of the Crab Nebula, and pulsation searches in the very high energy γ -ray range.

The Crab Nebula is known to be the most important calibration source for high energy astrophysics, and especially for the ground-based atmospheric Cerenkov technique (IACT). Since the LONS hereafter influences the trigger threshold, and hence energy and spectral determination of γ -rays, the LONS around the Crab Nebula must be accurately known in the wavelength interval corresponding to typical photomultiplier tubes used for the IACT. For instance,

3.1 The Crab optical spectrum and expected single photoelectron response. 73

Mirzoyan et al. [68, Mirzoyan et al., 1994] measured the LONS for La Palma over a large solid angle, obtaining a value of $I_{LONS} = (1.7 \pm 0.4) \times 10^{12} \text{ph} \cdot \text{m}^{-2} \cdot \text{s}^{-1} \cdot \text{sr}^{-1}$, at the site of the HEGRA experiment. In this chapter we will investigate the validity of this number for the Crab environment using the same photomultiplier tubes (PMTs) as used for the detection of Cerenkov showers. It also remains to be shown if different PMTs provide larger variations in the LONS compared to the variations expected across the angular size covered by the imaging camera.

The theoretical results, calculated before the observations to predict the feasibility of the observation, are discussed in the next section.

3.1 The Crab optical spectrum and expected single photoelectron response.

The optical spectra of the Crab pulsar main pulse and the spectrum of the underlying nebula [13, Carramiñana et al., 2000] are well known and follow the power laws given in table 1.

Flux	K	α
Main pulse	5.9×10^{-15}	$+0.2 \pm 0.1$
Interpulse	1.9×10^{-15}	$+0.2 \pm 0.1$
Underlying nebula	8×10^{-12}	-0.4 ± 0.1

Table 3.1: *Flux in the 0.1 phase window centered on the main pulse and interpulse. Where, $F_\lambda = K \times (\lambda/\lambda_o)^{-\alpha-2}$, $\lambda_o = 6000 \text{ \AA}$, $5000 \text{ \AA} < \lambda < 7500 \text{ \AA}$. Fluxes are in $\text{ergs} \cdot \text{s}^{-1} \cdot \text{\AA}^{-1} \cdot \text{cm}^{-2}$. The useful wavelength range is 3000-8000 \AA.*

In our study not only the nebula and pulsar spectra are needed, but we also have to take into account the influence of the LONS. Its intensity has been measured at La Palma, at the site of the HEGRA experiment.

$$I_{LONS} = (1.7 \pm 0.4) \cdot 10^{12} ph \cdot m^{-2} \cdot s^{-1} \cdot sr^{-1} \quad (3.1)$$

This is the intensity for a narrow angle detector (like CT1) in zenith direction, in the wavelength range of 300 to 600 nm at a mid-latitude site, when there are no bright stars in the field of view.

Using this expected spectral behavior, we can calculate the single photoelectron rate for CT1 and MAGIC. Note that the poor angular resolution (a single pixel with a FOV of 0.25° for CT1 and 0.1° for MAGIC), is compensated for by the large mirror area A (10 m^2 for CT1, 230 m^2 for MAGIC). The expected photoelectron rates are calculated with the following expression:

$$R = \int F_\lambda \cdot \eta \cdot QE_\nu \cdot A \cdot e^{-\tau_\nu} \beta \, d\nu, \quad (3.2)$$

where η is the mirror reflectivity and QE_ν the frequency dependent quantum efficiency of the PMs. The term $e^{-\tau_\nu}$ includes the atmospheric extinction. The parameter β takes into account two more terms, the photoelectron detection on the first dynode (0.9) and the light guide efficiency (90%). For CT1, a mean quantum efficiency QE of 20 % was used for the central photomultiplier. However to calculate the expected rates in the MAGIC central pixel, we used the quantum efficiency corresponding to a Hybrid Photo Detector (HPD) [38, Ferenc, 1997], which was proposed for the central pixel [62, Lorenz, 2003]. The QE of this kind of photo detector reaches values of $\sim 40 \%$ and the response of the photo detector, as a function of the wavelength, was used when calculating the rates.

Hence, for the Crab pulsar, the theoretical expected rates are given in Table 3.2, for MAGIC and CT1 telescopes.

If we assume at first that the PM has no inherent fluctuations (for example the

Telescope	Total Pulsar Rate (phe $\cdot$ s $^{-1}$)	Nebula Rate (phe $\cdot$ s $^{-1}$)	LONS Rate (phe $\cdot$ s $^{-1}$)	Background Rate (phe $\cdot$ s $^{-1}$)
CT1	1.505×10^5	1.410×10^8	3.294×10^7	1.740×10^8
MAGIC	3.742×10^6	3.506×10^9	1.311×10^8	3.638×10^9

Table 3.2: *Expected currents (in photoelectrons/s) induced by the optical emission of the Crab pulsar for the central pixel of the CT1 and MAGIC telescopes.*

fluctuations introduced by the electron multiplication process between dynodes, combined with the knowledge that $QE_\nu < 1$), we may derive the basic scaling parameter $x = p\sqrt{N}$ (see Eq. 2.30), for all test statistics, which measures the integrated difference between the measured pulse profile and a uniform distribution (no pulsed signal) [26, de Jager et al., 2001] in terms of significance. The parameter p is again the number of pulsed events, divided by the total number of events N (signal plus noise). This x also corresponds approximately to the expected DC excess (measured in Gaussian standard deviations). Assuming that the ephemeris for folding the arrival times of optical photons from the Crab pulsar to phases is correct, the expected scaling parameters for CT1 and MAGIC (assuming the Mirzoyan et al. [68] LONS is valid for a galactic source, although we expect it to be higher around the Crab pulsar) are:

$$X_{ct1} = 11\sqrt{\frac{T}{1sec}} \quad (3.3)$$

$$X_{MAGIC} = 62\sqrt{\frac{T}{1sec}} \quad (3.4)$$

Here it is assumed that the fluctuations in N are Poissonian and that the noise contribution to N is purely due to the LONS. From the basic scaling parameter we can derive the expected time to detect the optical pulsation with a given

significance. The theoretical expected rate from the Crab pulsar shows that it would be possible to detect the Crab optical pulse with the CT1 telescope in $T=75$ ms, with a significance of 3σ , and even better with MAGIC where the main pulse can be detected within 5 ms, which is the approximate duration of the pulse. This implies that MAGIC, theoretically and using the value of the LONS intensity given by Mirzoyan et al. [68], should have the capability of monitoring the Crab on a pulse-by-pulse basis based on theoretical arguments, and using the LONS intensity corresponding to an extragalactic source.

This assertion is only valid if the photo detector (with 100% quantum efficiency) itself has no noise and the only source of fluctuations is Poissonian in nature, due to the signal. The expected detection time increases dramatically when we take fluctuations into account, as shown by observations presented below.

Due to the geographical latitude of MAGIC, there are only two other candidates detectable by MAGIC, since the intensity of their optical emissions are too faint for CT1. MAGIC, due to the huge reflector, could have a chance to detect Geminga and PSR B0656+14.

The Geminga (PSR B0630+18) spectrum shows a continuous power law $F_\nu \sim \nu^{-0.8 \pm 0.5}$ [42, Gil et al., 2001] from 3700 to 8000 Å, most likely due to the dominant nonthermal synchrotron component. The magnitude of the pulsed emission is $m_B = 26.0 \pm 0.4$ and it is found that there is a phase agreement of the optical data with γ -ray and hard X-ray curves. Normalizing this expression to the blue band ($\lambda_o = 4400$ Å) the optical spectrum of Geminga can be defined:

$$F_\lambda = K \left(\frac{\lambda}{\lambda_o} \right)^{-\alpha-2} \text{ erg cm}^{-2} \text{ s}^{-1} \text{ Å}^{-1} \quad (3.5)$$

where $\alpha = -0.8 \pm 0.5$, and $K = 2.7 \times 10^{-19}$. Hence, from the rate values

found for Geminga (~ 124 Hz) we should be able to detect a very faint source like Geminga within less than 24 hours with 3σ significance, since these calculations have been done assuming a HPD detector in the MAGIC central pixel and therefore a higher sensitivity in the blue wavelength.

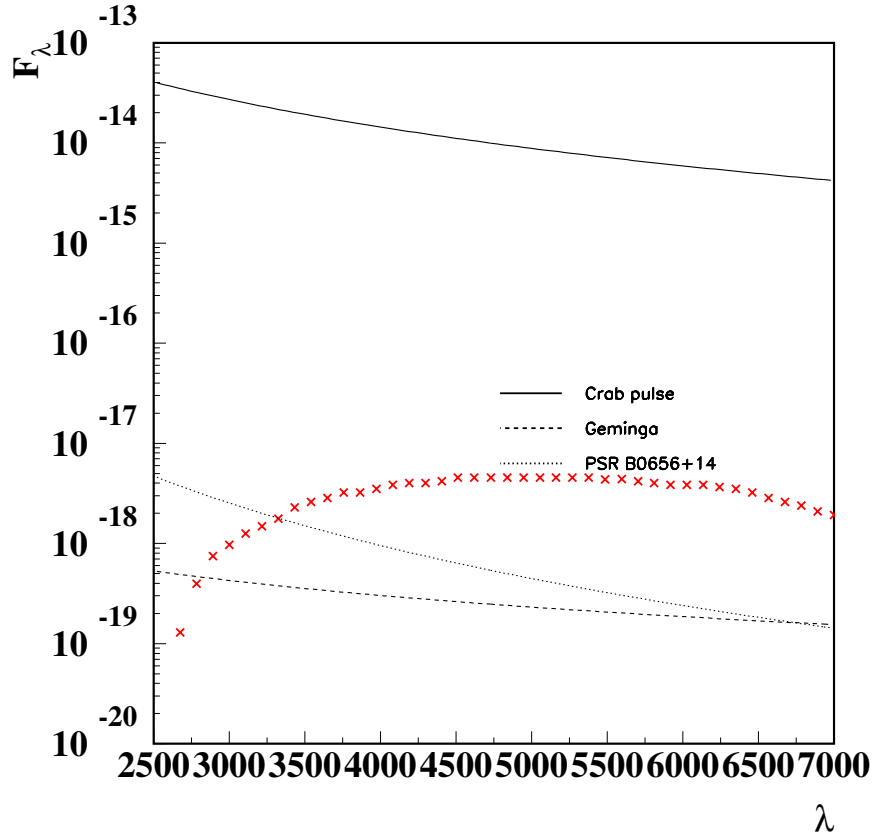


Figure 3.1: Optical pulsed spectra F_λ ($\text{erg cm}^{-2} \text{s}^{-1} \text{\AA}^{-1}$) for pulsars detectable with MAGIC: Crab, Geminga and PSR B0656+14 as a function of wavelength. The red crosses mark the quantum efficiency for an HPD, multiplied by 10^{-19} . It is clear that the detector will be more sensitive at lower wavelength with these type of photodetectors.

The other optical pulsar candidate is PSR B0656+14. Its optical spectrum also follows a power law as a result of synchrotron radiation (in this case $K = 6.9 \times 10^{-19}$

and $\alpha = 1.4^{+0.6}_{-0.7}$). This pulsar is brighter than Geminga with magnitude $m_B = 25$. Its optical radiation is more blue (due to the higher spectral index), which makes it easier to detect by MAGIC.

With an expected pulse rate of 360 Hz and regarding only the LONS as background, the detection in this case with 3σ will be possible in only ~ 4 hours.

Figure 3.1 shows the optical spectra for Crab, Geminga and PSR B0656+14. The last two have very similar fluxes although the slopes are different, making PSR B0656+14 emission dominant at lower wavelengths where MAGIC central pixel will be more sensitive. The response of the HPD (quantum efficiency multiplied by 10^{-19}) is plotted as well in the Figure, showing higher sensitivity at lower wavelengths [12, Camara, 2003].

Therefore, among the known optical pulsars, the next candidate to consider after Crab will be PSR B0656+14, if such long observation times are allowed.

The spectrum from this pulsar is however stepper compared to Geminga [82, Ramanamurthy et al., 1996] so Geminga is favored from the viewpoint of γ -ray observations above 1 GeV.

3.2 Experimental Setup

3.2.1 First Test in CT1

The first tests were performed in La Palma in February 2002, with an attempt to detect the pulsed optical signal of the Crab pulsar with CT1. The tests we did are described in this section.

We made use of the PM already installed on the central pixel of the CT1 camera. Monitoring the DC current while pointing at the Crab pulsar, we expected to detect an increase in it due to the optical pulsation. The output signal of a PM in CT1 is split in two branches: a “fast or **AC branch**” designed to transmit the fast signals generated by the showers (on the order of some ns) and a “slow or

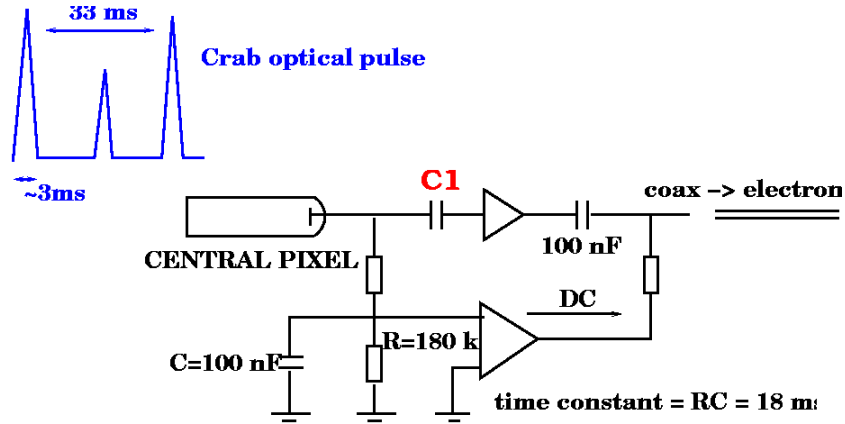


Figure 3.2: *Exchange in the base of the central pixel to reduce the integration time*

DC branch” designed to monitor the DC current of the pixel. Figure 3.2 shows a scheme of the electronic configuration of the PMs, which is very similar to the one installed in the MAGIC camera. Both branches are transmitted over the same coaxial cable to the electronics room. The slow branch has a long integration time of the order of 18 ms given by a 180 K Ω resistor and a 100 nF capacitor.

For the purposes of our experiment this time constant must be smaller than the typical time of the pulse variation. Taking into account that we want to detect a pulse of ~ 3 ms, we replace the base with a new one. This new base works with a resistance of $R = 2$ M Ω and a capacitor $C = 100$ pF. The new time constant will be 0.2 ms. For the data acquisition we use the card PCI NI6034E of National Instruments, and the software was modified to record the current every 0.5 ms for 2 hours.

Finally, a high pass filter of $\tau \sim 100$ ms was installed to reduce the high background due to the LONS.

Since no indication of the Crab pulsar was found after modifying the base, we installed a green LED in the plate of the telescope with the aim of flashing the telescope camera to calibrate the system, with pulses of ms length and digitize

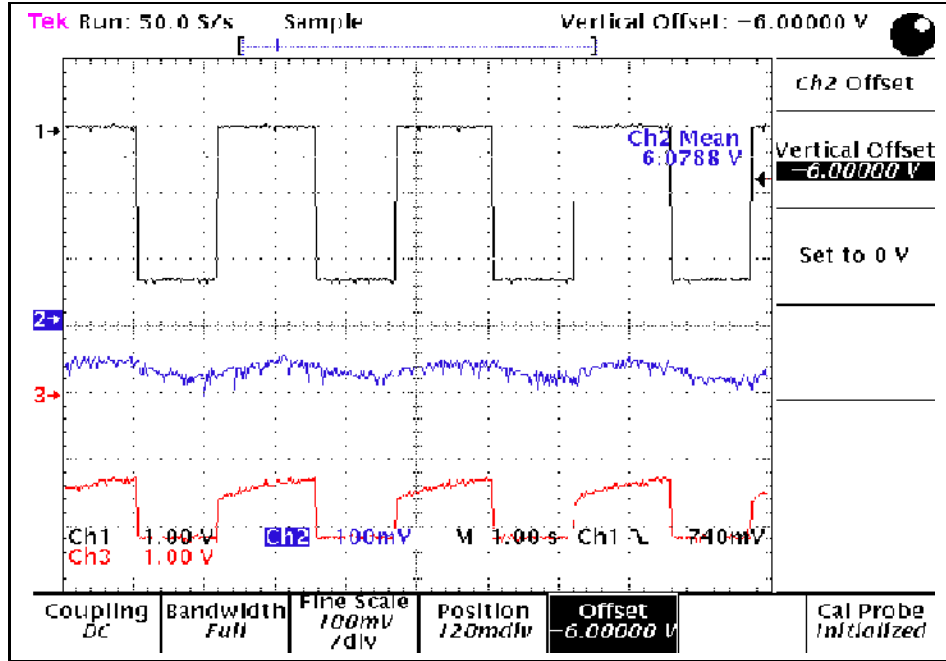


Figure 3.3: Channel 1 (in black) is the voltage signal used to drive the LED ($50\ \Omega$ termination). Channel 2 (in green) is connected to the base of the central pixel ($1\ \text{M}\Omega$ termination) and shows rise and fall times of the order of half a second. Channel 3 (in red) is again the line going to the LED, with $1\ \text{M}\Omega$ termination.

them using the card. This LED was driven by a NIM pulse generator.

After a few tests we realized that the fast light pulses resulted in much slower pulses in the pixel current. Fig. 3.3 shows the pixel current pulse registered with a fast oscilloscope at the pixel base. The fall time is of the order of $0.5\ \text{s}$, which is obviously not appropriate for the search for fast optical pulses from pulsars. The first guess was that this green LED had a very slow reaction time and that the pixel base was still not fast enough or that some element in the transmission line was widening the signal.

On the other hand, we suspected that the real value of the LONS was higher than the theoretically assumed, since no excess over the LONS due to the nebula flux was detected.

3.2.2 Second Tests and Final Observations

Before continuing with test on CT1, further tests at IFAE Barcelona were performed, in July 2002, using just a PM with an appropriate base.

The aim was to understand the problem as discussed in Section 3.2.1, resulting in the construction of a more optimal setup. The Crab optical pulse was simulated using a fast LED.

Again, to be able to detect the 33 ms from the Crab optical pulse simulation, we altered the PM base as discussed below. In order to reduce the integration time we replaced the original capacitor of 100 nF in the DC branch with a new one of 1 nF.

To simulate the Crab pulsations, a fast LED, (Nichia type NSPB 300 AS, of 265 nm wavelength and 4 mW output) was used, but located 7.5 m away from the PM. The luminosity given by the LED at this distance is $\sim 1 \times 10^{11} \text{ ph} \cdot \text{s}^{-1}$, while for the Crab pulsar we receive theoretically $\sim 2 \times 10^7 \text{ ph} \cdot \text{s}^{-1}$. To reduce the LED to levels comparable to the Crab pulsar, a filter was installed. The reduction factor of this filter was of $\sim 10^3 - 10^4$ at the LED wavelength. The light pulse was attenuated to a level comparable to that of the LONS at La Palma, but still higher than the expected Crab pulse and in a range where we could already attempt realistic tests. The PM DC current was taken directly from the camera mother board. The output signal was analyzed not only with the digitization card, but also with an oscilloscope.

After several tests, we realized that the rise and fall times of the DC current pulses was still too slow.

Figure 3.4 shows a fall time ~ 20 ms as detected with the digitizing card when the LED is fired for very short square pulses. Only then we realized that the capacitor in the AC branch was still delaying the signal. The rise and fall times of the DC current in the PM could be made fast enough by adjusting the capacitor and resistance, shown in Fig. 3.2 to a value such that RC time was

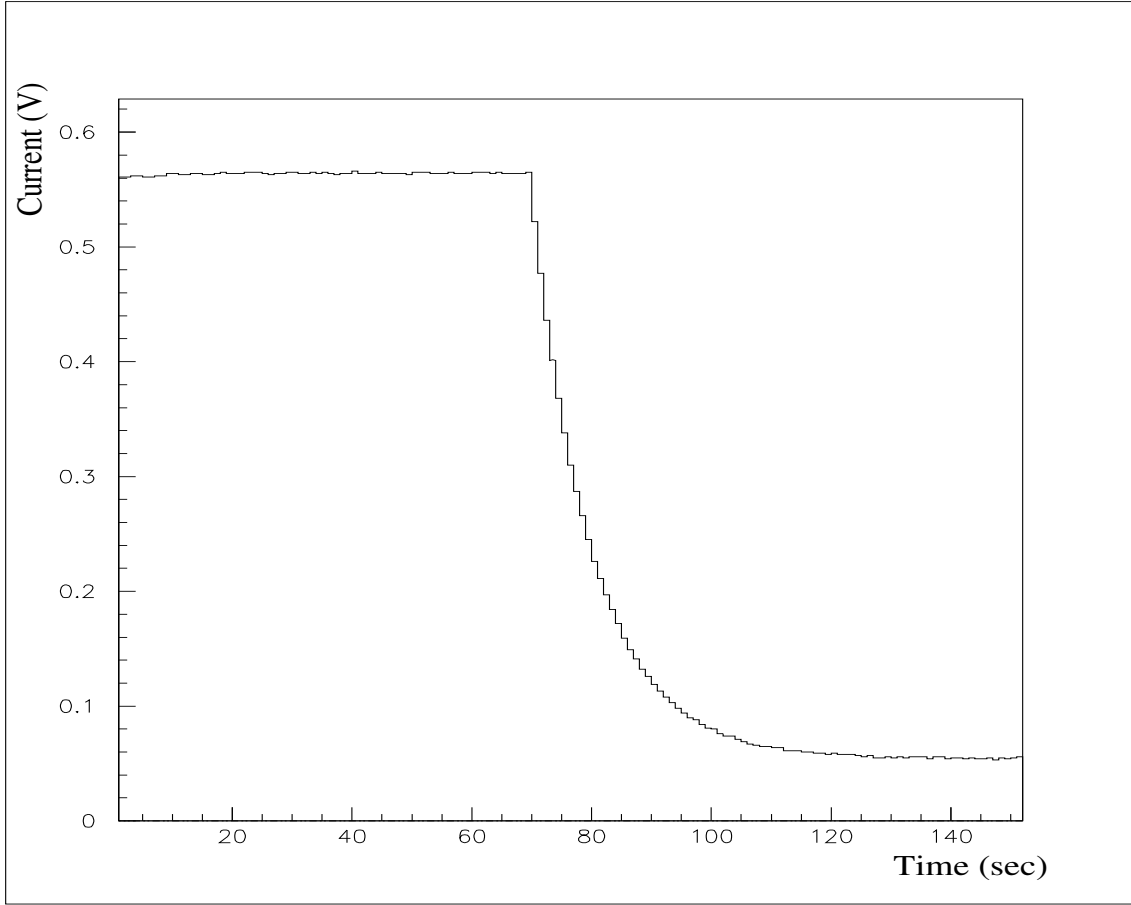


Figure 3.4: *Fall time measured with the digitization card using a fast LED but too slow a PM base (0.5 sec).*

smaller than the duty cycle of the Crab pulses.

Figure 3.5 shows the digitized signal in the card for a run of 200 samples at 1000 Hz and in the oscilloscope. Rise and fall times were measured to be shorter than $10 \mu\text{s}$.

The final observation were taken on November 1-8th 2002. As discussed above, the PM #1 was modified. The **DC branch** was then changed to be able to detect the Crab pulse, and the **AC branch** had to be removed [77, Oña-Wilhelmi et al., 2002]. The pixel line was AC coupled to reject low frequencies and increase the digitizing card effective dynamic range.

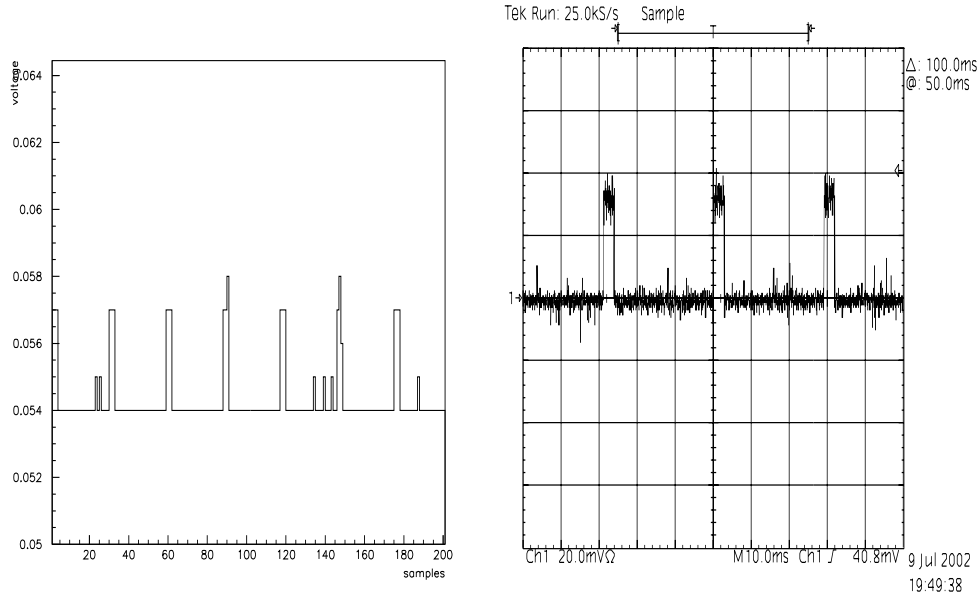


Figure 3.5: *Detection of the simulated Crab optical pulsed signal with the digitization card (left) and the oscilloscope (right). Each time sample in the card corresponds to 1 ms. Voltage is measured in volts.*

Datasets were obtained with the Crab centered in the central pixel, with OFF source runs to isolate systematic effects.

3.3 Analysis and Results of the Pulsed Optical Signal

The observed arrival times t_i were folded to obtain the phases Φ_i .

$$\Phi_i = \Phi_0 + \nu(t_i - t_0) + \frac{1}{2}\dot{\nu}(t_i - t_0)^2 \quad (3.6)$$

where $\nu, \dot{\nu}$ are the Crab pulsed frequency and frequency derivative, and t_0 is the reference epoch (in Modified Julian Day (MJD)).

To transform these arrival times to the solar barycentre system, we used the TEMPO software [65, Manchester et al., 2001]. A correction due to a small average drift in the digitisation card clock, as measured with respect to an atomic

reference clock, was also introduced. The Jodrell Bank ephemeris (frequency, first frequency derivative and reference time) applicable to these observations are given in Table 3.3 (Jodrell bank Crab pulsar monthly ephemeris) ([http : //www.jb.man.ac.uk/ ~ pulsar/crab.html](http://www.jb.man.ac.uk/~pulsar/crab.html)).

Date	T_o (MJD)	ν (Hz)	$\dot{\nu}$ ($10^{-15} \text{ sec}^{-2}$)	ΔT (sec)
15th October 2002	52562	29.8132421622	-373777.06	0.024215
15th November 2002	52593	29.8122410972	-373758.90	0.018591

Table 3.3: *Radio Timing Parameters for the Crab pulsar.*

A phaseogram was produced for each independent frequency confined to the wide frequency range between 26.0 to 32.0 Hz. The independent Fourier spacing (IFS) between such independent frequencies is known to be $1/T$, where T hours is the observation time (2 hours). The folded intensities for each test frequency were tested against a uniform distribution by performing a χ^2 fit to a constant intensity. The statistic χ^2 applied to current is defined like:

$$\chi^2_{k-1} = \sum_{i=1}^K \left(\frac{x_i - \mu}{\sigma_i} \right)^2 \quad (3.7)$$

where x_i is the DC current values, μ the mean value and σ the standard deviation.

The reduced χ^2 was calculated (Figure 3.6). A maximum reduced χ^2 of ~ 154 was found and the periodogram was found to peak within one IFS from the expected Jodrell bank ephemeris, consistent with the uncertainty in the drift expected from the clock on the digitation card. Absolute phase matching with the expected radio/optical/X-ray/ γ -ray main pulse position was not possible, since the absolute UTC match was not done and the drift was not accurately measured at the time of observations.

Furthermore, the frequency resolution for the 2 hour observation was also

too coarse to resolve the differences between the ephemerides corresponding to 15 October and 15 November. The predicted frequency shown in Figure 3.6 is therefore consistent with both ephemerides.

Figure 3.6 shows the well-known Crab-like pulse profile, with two sharp peaks, and the smaller inter pulse ~ 0.42 in phase after the dominant main pulse, as expected. The corresponding pulsed fraction can now be used to isolate the LONS and cross-check the measurement made by Mirzoyan et al. [68] Using that empirical correction we can infer the real situation which we expect to have in MAGIC.

There was a relatively small glitch the day after the last observation, between MJD 52587.2 and 52587.3, but unfortunately we were not able to measure it since the change in frequency was smaller than the card precision. The setup in MAGIC central pixel will be controlled by an atomic clock in the future, so that the high timing precision will allow a measurement of any discontinuity in frequency.

The corresponding pulsed fraction can now be used to isolate the LONS to verify the measurement made by Mirzoyan et al. [68]. This new measured LONS may be used for future MAGIC observations and calibration associated with the Crab region.

3.4 Determination of the LONS around the Crab pulsar

In this section we will determine the galactic Night Sky Background (LONS) around the Crab pulsar by using the pulsed fraction of the signal detected in CT1. A realistic result for the LONS intensity is necessary since the Crab pulsar is the standard candle in γ -ray astrophysics and is also used to calibrate the telescope. This result will be checked by comparing with the LONS calculated from the intensity of the CT1 camera response to ζ Tauri.

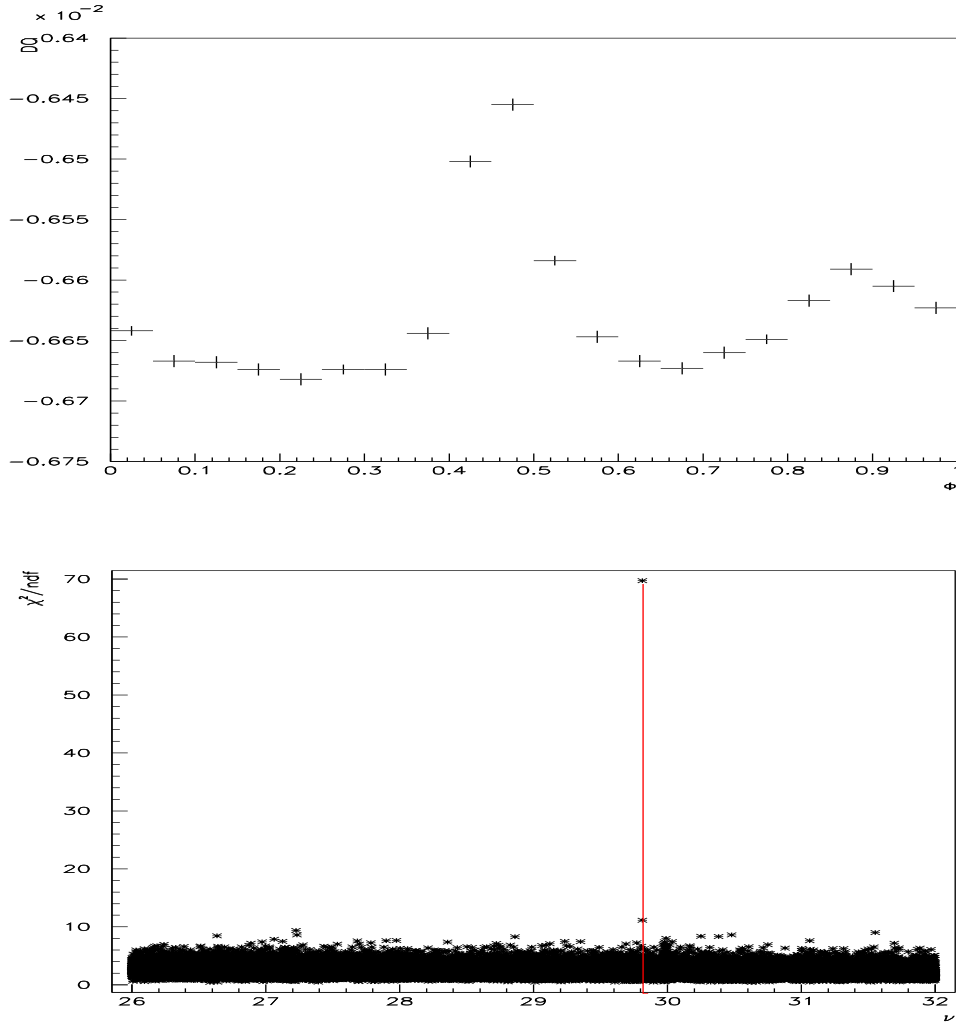


Figure 3.6: *Observations of the optical Crab pulsar with CT1. On the upper chart (a) the two peaks are clearly visible in the light curve with a separation in phase of $\sim 0.4^\circ$. On the lower chart (b) the excess in the reduced χ^2 searching in a wide range of frequencies, using the 15th October ephemeris but keeping the frequency ν variable. The difference between the frequency found with 15th October ephemeris and 15th November is consistent with the drift in the digitization card clock. The figure on the bottom shows a zoom out of the plot on the right, the value of the reduced χ^2 vs. frequency, using ephemeris of 15th October.*

The LONS has several components, which include the zodiacal light, integrated starlight, air-glow, aurorae, diffuse galactic light and integrated cosmic light. In

case of the Crab pulsar the main effect is the integrated starlight because it is on the galactic plane. The zodiacal component is very weak or negligible since with atmospheric Cerenkov detectors one performs measurements during dark nights (between two astronomical twilights in the evening and in the morning). Also the aurorae are rare phenomena at mid latitude places.

Thus, to calculate the LONS from the optical pulsed detection of Crab, we should introduce the concept of the density function of the light curve or periodic shapes:

In the pre-analysis discussed in the previous section, we folded (or superimposes) the arrival times on the phase interval $(0,1)$ (or $(0,2\pi)$) using the appropriate parameters (e.g. period, period derivative etc.) of the periodic source. This interval represents one full period of rotation and in the absence of any periodicity, the folded events will be uniformly distributed if the frequency is larger than the mean count rate (i.e. "sparse data"), and one can test for periodicity by applying a *test for uniformity* on the circle. We used the Pearson's χ^2 -test with K bins. To test for the presence of a periodic signal one should actually test for the absence of such a signal (i.e. the null hypothesis H_o), which is usually characterized by a test of uniformity. Let $f(\theta)$ be the periodic density function or light curve (empirically known) of the folded arrival times. The null hypothesis may then be written as follows

$$H_o : f(\theta) = \frac{1}{2\pi} \quad (3.8)$$

with $\theta \in (0, 2\pi)$.

In the presence of the periodic signal $f(\theta)$ will differ from Eq. 3.8 and may be written as follows under the alternative hypothesis

$$H_A : f(\theta) = pf_s(\theta) + (1-p)/2\pi \quad (3.9)$$

where p is the strength of the periodic signal. The source function $f_s(\theta)$ gives the relative radiation intensity as a function of θ , and the density function is properly normalized.

The off-pulse interval (region of minimum intensity) following the inter pulse and preceding the main pulse was used to estimate the pulsed fraction. Taking the errors in each bin properly into account (exploiting the Central Limit Theorem) we derive a pulsed fraction for the Crab of $p = (5.00 \pm 0.02) \times 10^{-4}$.

To check the stability of the signal with time, the total data set of two hours was divided into six data sets of 20 minutes each. The pulsed fraction, as well as the ratio between the first peak and the second peak was calculated for each subset. An indirect indication of the stable tracking of the telescope is that the pulsed fraction would have changed if the telescope pointing changed by even less than a pixel size, but no evidence of this can be seen during the two hour observation time window. The pulse ratio should remain stable if the test frequency and true pulsar frequency are similar within statistics. The results are shown in the figure 3.7. It is clear that both parameters appear to be stable within statistical errors.

Once we have calculated the pulsed fraction, we can compare it with the theoretical value given in the previous section, which is equal to $p_{theor} = 8.6 \times 10^{-4}$. The discrepancy is attributed to the uncertainty on the measurement of the LONS, the influence of the mirror point spread function PSF, and the fact that the Crab pulsar was off-center due to uncorrected but constant miss-pointing in the telescope.

To determine the real LONS around the Crab pulsar, we first need to know the point spread function of the telescope. To this end, we used a point source to measure the amount of light which falls within the central pixel, as well as the miss-pointing of the telescope.

The selected source was ζ Tauri (ICRS 2000.0 coordinates are $5^h:37':38.70''$ RA, $21^\circ:08':33.0''$ DEC with a magnitude of 2.84), which is in the same CT1 total

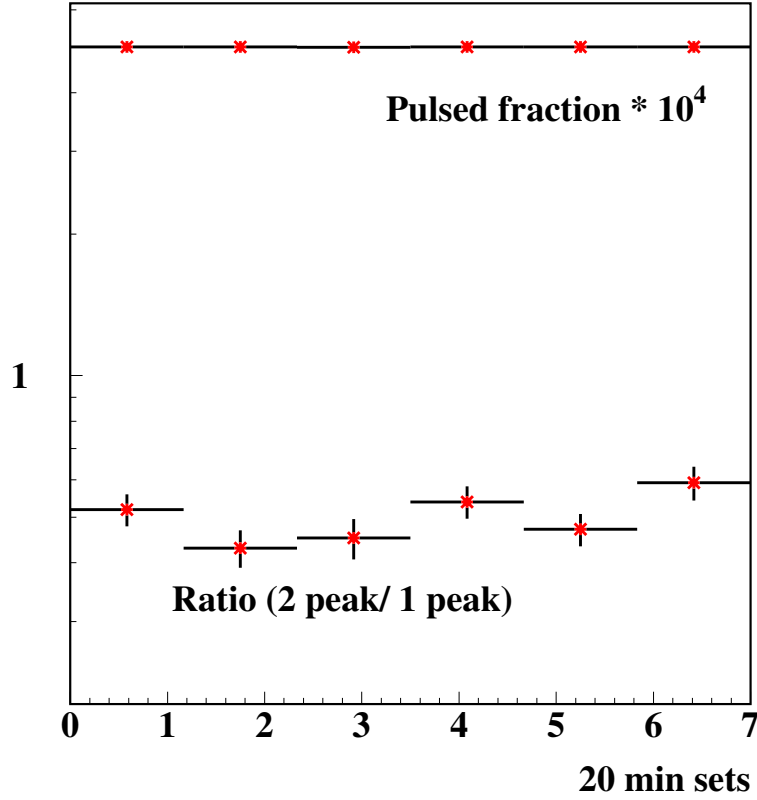


Figure 3.7: *Test of the telescope response to the pulsed signal in two hours. In the upper chart (a) of the plot it is shown the pulsed fraction times 10^4 for the six 20 min set. The ratio between the first and second peak remains constant with time, represented on the lower chart.*

camera field of view as the Crab pulsar. The advantage in using this reference star is that we know its photometric properties and the atmospheric extinction for both the pulsar and star will be the same, given their proximity. The PSF of the telescope was determined from a raster scan with a resolution finer than the size of a pixel (see Fig. 3.8(a)), with zero position corresponding to the expected position of the star. Since the telescope is focusing at ~ 6 km (where the Cerenkov shower has its maximum), the image of a point-like source will be smeared out around 0.14° .

The signal from ζ Tauri is slightly shifted to lower right ascension values,

due to the miss-pointing in the telescope. This was rectified by calculating the maximum value of the signal, and measuring the difference between it and the true value. A schematic figure of the CT1 central pixel is also drawn. To calculate the true signal collected by the central pixel, the signal intensity was fitted to a bidimensional Gaussian (fig. 3.8 (b, c)), and the fraction of the signal within the central pixel was measured to be:

$$\Pi = \frac{\int_0^{r_{pixel}} \int_0^{2\pi} G(r, \theta) dr d\theta}{\int_0^{\infty} \int_0^{2\pi} G(r, \theta) dr d\theta} = 69\% \quad (3.10)$$

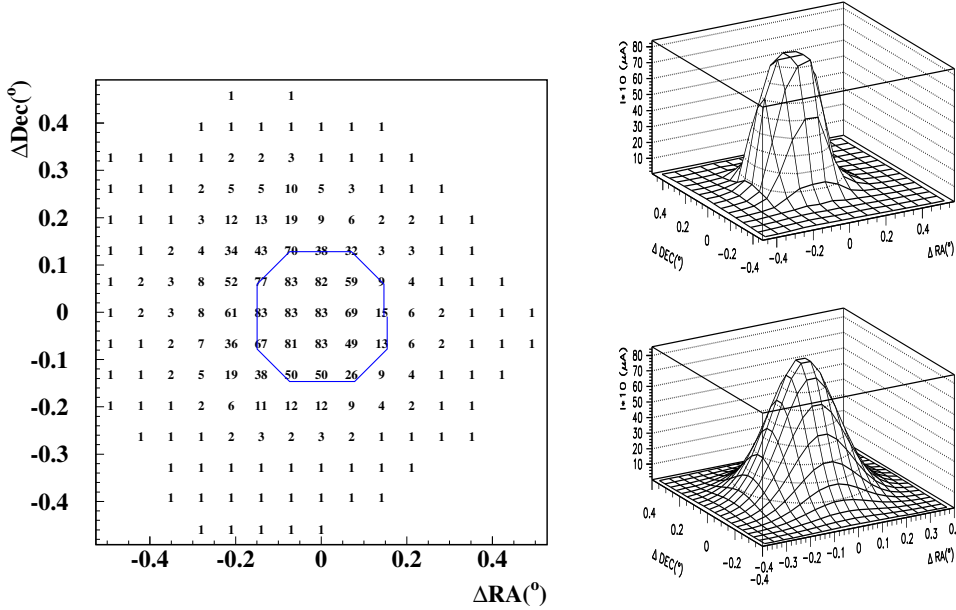


Figure 3.8: *DC currents induced by ζ Tauri on the CT1 central pixel, in $0.1 \mu A$ units. To the left, the signal of ζ Tauri in the CT camera is shown, where the background is already subtracted and the position is relative to the real position of the star, in right ascension and declination (where the maximum of intensity is). The approximate shape of the central pixel is drawn to illustrate how the signal is spread due to the large PSF. Top right (b) ζ Tauri intensity in a 2-dimension plot, is shown compared with the same signal fitted with a bidimensional Gaussian on the bottom.*

Once we calculated the correction $\Pi(x_o, y_o, \delta x, \delta y)$ for the miss-pointing, it was possible to derive the LONS rate from:

$$p = \frac{R_{pulsed}}{R_{pulsed} + R_{nebula} + R_{LONS}} \quad (3.11)$$

where p is the empirical pulsed fraction, R_{pulsed} is the expected rate of pulsed events, R_{nebula} the expected rate coming from the Crab nebula and R_{LONS} is the galactic background around the Crab nebula to be determined.

Hence, assuming that the contribution of the pulsed signal is negligible compared to the background due to the nebula and the LONS, we have:

$$p = \frac{\int F_{\nu}^{pulsed} \cdot \eta \cdot QE \cdot A_{ct1} \epsilon_{\gamma v} e^{-\tau_{\nu}} \Pi d\nu}{\int F_{\nu}^{nebula} \cdot \eta \cdot QE \cdot A_{ct1} \epsilon_{\gamma v} e^{-\tau_{\nu}} \Pi d\nu + \int F_{\nu}^{LONS} \cdot \Omega \cdot \eta \cdot QE \cdot A_{ct1} \epsilon_{\gamma v} d\nu}, \quad (3.12)$$

where η is the mirror reflectivity, Q the quantum efficiency of the PMs, A_{ct1} is the CT1 mirror area, Π is the fraction of light from a point source captured by the central pixel, $e^{-\tau_{\nu}}$ takes into account the atmospheric extinction at La Palma, Ω is the solid angle subtended by the central pixel, and $\epsilon_{\gamma v}$ is a conversion factor between number of photoelectrons and voltage measured by the digitisation card, which cancels in the above expression.

Therefore, using the pulsed fraction of the total Crab signal, the frequency integrated LONS (weighted with the frequency dependency of the quantum efficiency of the PM), measured as an intensity, is determined from Eq. 3.12 as:

$$I_{LONS} = \int F_{\nu}^{LONS} d\nu = 4.6 \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \cdot \text{s}^{-1} \cdot \text{sr}^{-1} \quad (3.13)$$

To cross-check this result we use the signal detected by CT1 from ζ Tauri

in the central pixel (see Fig. 3.8). If i_{cp} is the total observed current intensity from the direction of ζ Tauri in the central pixel alone, $i_{\zeta Tauri}$ is the expected contribution from ζ Tauri alone to the current in this central pixel, and $\epsilon_{\gamma i}$ is the conversion factor from photoelectrons to current (in μA), the conversion factor can be calculated from:

$$\epsilon_{\gamma i} \Pi \int F_{\nu}^{\zeta Tauri} \cdot \eta \cdot Q \cdot A_{ct1} d\nu = i_{\zeta Tauri} = i_{cp} - i_{LONS} \sim i_{cp}. \quad (3.14)$$

The last approximation follows from the knowledge that the central pixel's current is dominated by this bright star rendering the contribution from the LONS in the central pixel negligible. Integrating the known spectrum of ζ Tauri as above (including the effect of atmospheric extinction), we obtain a value of $\epsilon_{\gamma i} = 2.95 \times 10^{-9}$ from observations. Using this information we can convert current to photoelectron rate, and by subtracting the PSF of ζ Tauri from a square degree image centered on this star (shown in Figure 3.8), we can obtain the LONS level. The error on the LONS is minimal near the outer wings of the PSF, since the LONS is dominant in that area. An annular region is used to estimate the LONS, giving

$$I_{LONS} = \int F_{\nu}^{LONS} d\nu \sim 4.3 \times 10^{12} ph \cdot m^{-2} \cdot s^{-1} \cdot sr^{-1}, \quad (3.15)$$

after normalizing to the solid angle subtended by the selected annular region.

The two values for I_{LONS} calculated by two independent methods agree within $\sim 7\%$. Both of them give us a value for the galactic LONS, which are 2.5 to 3 times higher than the measurement of Mirzoyan et al. [68], used for the theoretical predictions, as expected. Moreover, the LONS intensity calculated from the pulsed fraction is more accurate, with the uncertainty on the correction

factor Π entering as the major uncertainty, whereas the LONS determination from ζ Tauri also includes the uncertainty in the definition of the background region around the star.

3.5 Sensitivity for Pulsed Detection in Optical

To determine the sensitivity of CT1 to detect the optical Crab pulsar and then be able to infer the future sensitivity of MAGIC, we may define a relative significance in terms of a "standard deviation" s . To determine whether a periodic signal is present in the data, one should test the hypothesis $H_o: p=0$ against $H_A: p>0$ and to do this we use a measure of the difference between $f(\theta)$ and the uniformity density $1/2\pi$. A good measure is given by the functional:

$$\psi(f) = \int_0^{2\pi} (f(\theta) - 1/2\pi)^2 d\theta, \quad (3.16)$$

Clearly, the hypothesis for uniformity should be rejected when $\psi(f)$ is too large. However, $\psi(f)$ is unknown (since $f(\theta)$ is unknown) and one therefore has to estimate this functional. This can be "crudely" done by replacing f in $\psi(f)$ by a consistent estimator $\hat{f}_h$, which depends on a data smoothing parameter h .

The statistical literature [27, de Jager et al., 1989] gives various choices of an estimator $\hat{f}_h$ to replace the functional ψ . In our case this estimator is a histogram, where $k = h$ gives the fixed corresponding number of bins for the histogram and Eq. 3.16 results within a constant factor to the well known χ^2 statistic of Pearson, which is asymptotically χ^2 distributed with $k-1$ degrees of freedom under H_o . In the case of counting statistics (e.g. if single photon counting can be achieved) the χ^2 test for uniformity will be given by:

$$\chi_{k-1}^2 = 2\pi n \psi(\hat{f}_k) = \sum_{j=1}^k \frac{X_j - n/K}{n/K} \quad (3.17)$$

where X_j is the number of events in the j th bin.

The sensitivity of the CT1 telescope to detect the Crab optical pulse can be expressed in terms of a relative "standard deviation" s of the signal relative to the behavior of reduced χ^2 at frequencies different from the Crab frequency:

$$s = \frac{\chi_{red}^2 - \mu_{background}}{\sigma_{background}} \quad (3.18)$$

for each observation time interval, where $\mu_{background}$ is the mean value of χ_{red}^2 at such offset frequencies and $\sigma_{background}$ is the corresponding standard deviation, which is constant for all selected times T .

The total observation was split into $L = 7200$ s/T independent datasets, each of length T , from which the average value of the statistic s was calculated for $T=50$ s to 7200 s.

It can be shown that this average value of s is:

$$s \sim p^2 \cdot T \cdot g(k, f_s), \quad (3.19)$$

where k is the number of bins, f_s the signal pulse shape and $g(k, f_s)$ a function depending on these quantities. Note that s is not a Gaussian standard deviation, since the number k is too small, but we expect that $\langle s \rangle \sim \sigma^2$, where $\sigma \propto p \cdot \sqrt{T}$ would be the Gaussian significance of a DC detection of the same signal. Figure 3.10 shows a graph for $\log \langle s \rangle$ vs. $\log T$, showing a linear relationship between s and T as predicted..

To prove Eq. 3.19, we consider a counting experiment with one of k bins

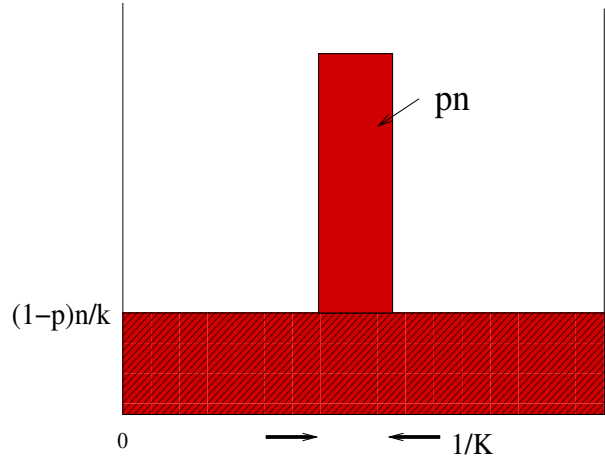


Figure 3.9: *Schematic plot of a signal in one bin to show the linear dependency of s with time*

containing pn excess counts as represented in Fig 3.9, whereas the background level in the rest of the $k-1$ bins is on average $(1-p)n/k$. It can be easily shown that the average χ^2 statistic would be given by:

$$\langle \chi_{k-1}^2 \rangle \approx (k-1) + p^2 n(k-1) \quad (3.20)$$

where the number $(k-1)$ represents the average value for χ_{k-1}^2 in the absence of a signal, whereas the term $p^2 n(k-1)$ represents the signal term, which is proportional to $p^2 T$ (Eq. 3.19), where T is the observation time and g is a normalising factor for the statistic Eq. 3.19 to compensate for the fact that the output (in terms of current) may not represent counting statistic, as assumed in the above derivation.

It is thus clear that the statistic s should increase lineally with observation time T .

Using the calculated value of the galactic LONS from the pulsation analysis, it is possible to estimate the time for MAGIC to detect the pulsar in optical wavelengths, given the much larger mirror area, improved quantum efficiency,

and pixel size (FOV of 0.1°), but similar non-Poissonian fluctuations in the PM.

By scaling the graph of $\log \langle s \rangle$ vs. $\log T$ accordingly, we can see from Figure 3.10 that a detection should be possible with 30 sec using MAGIC.

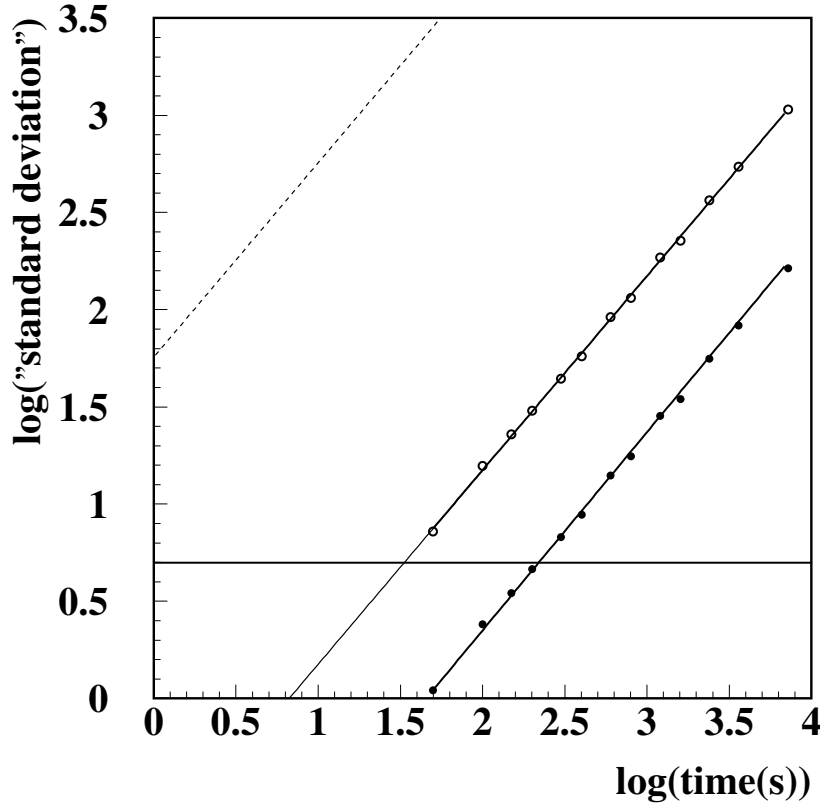


Figure 3.10: *log of "Standard deviation" of the signal related to the background versus observation time. The CT1 empirical data are represented with filled bullet and the empty circles represent the extrapolation for the MAGIC telescope. The horizontal line shows the limit at $s=5\sigma$. The dashed line represent x^2 for an HPD and single pulse detection for Crab.*

The new value of the scaling parameter x for MAGIC, using a corresponding galactic LONS and a high quantum efficiency HPD, will be $x = 57\sqrt{T}$. In Fig 3.10 the x^2 is also represented. The dashed line then represents the MAGIC sensitivity for an HPD with 45% QE, for which the detector background is negligible and the counts in the HPD are approximately Poissonian.

3.6 Further Applications in MAGIC

The detection of the Crab optical pulsation with CT1 provides us with the opportunity to learn from the experience and develop an improved system for MAGIC. With these improvements the detection of the Crab will be done in ~ 7.7 ms, but also the MAGIC central pixel will be possible to use as a Optical Camera, performing gamma and optical contemporary observations.

Other improvements that can be done in MAGIC:

- Since the telescope is focusing at 10 Km from the telescope, the image of a point-like source will increase from $d50 = 0.5$ cm to $d50 = 2.2$ cm, where $d50$ indicates the diameter of a circle containing 50% of the reflected light [7, Bigongiari et al., 2003]. Applying a proper lense as an ocular at the entrance of the central pixel, we will be able to refocus the reflected light, collecting more light within the pixel (in red) (see Fig. 3.11).
- High quality PMs can be used instead the normal PM used in the whole MAGIC camera. An array of HPD, with quantum efficiency of $\sim 40\%$ will improve the signal by a factor of ~ 1.4 .

3.7 Conclusions

The detection of the Crab optical pulsation was performed with the HEGRA CT1 telescope, which is possibly the smallest Cerenkov telescope to have seen the Crab optical pulsations. A detection required an observation time of 200 seconds. Since the design of the CT1 PM base is very similar to that of MAGIC we conclude that the detection with this new generation Cerenkov telescope is possible and should be achieved in a substantially shorter time than for CT1. From the CT1 results we have shown that a detection can be achieved within at least 30 s by the increasing the currents. However, using a single photon counting technique,

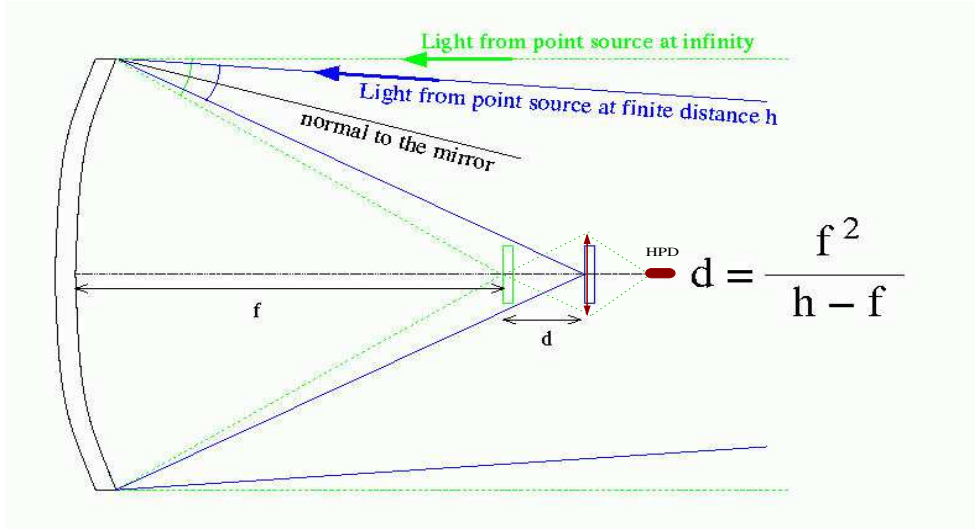


Figure 3.11: *Schematic design of the ocular lens in the central pixel to refocus the light from a point-like source. The calculated distance d is 3 cm. The PM then has to be shifted back to collect the refocused light, with the shift depending on the type of lense used.*

facilitated by the use of high quantum efficiency photo detectors (HPDs), it has been proved (Fig. 3.10) that the Crab pulse may be detected in 7 ms, and we may even be able to resolve single pulses if some improvements are applied.

The use of a PM base with the technical modifications described above is, however, incompatible with normal VHE observations. The modification could only be considered for the central pixel of the telescope camera. Optical data on the Crab pulsation taken simultaneously with regular VHE observations could be very important for monitoring atmospheric extinction. In fact, high quality observations should show a very specific zenith angle dependency of the pulsed fraction. Thus, the "seeing condition" at a given time may also be expressed in terms of the observed pulsed fraction corrected for atmospheric extinction. Furthermore, apart from using optical observations to provide a timing ephemeris for VHE observations, the timing analysis software can also be tested using optical observations.

Frequent monitoring of the Crab by MAGIC (for calibration purposes) should also improve the chance of detecting glitches. In fact, very specific software can be written to test for glitching behavior during observations. Such a specially designed central pixel detector with high quantum efficiency, and relatively low detector noise would be very useful for monitoring the state of optical emission from several candidate γ -ray sources such as cataclysmic variables, X-ray binaries, microquasars, AGNs and possibly GRBs.

The galactic light of the night sky rate on the surroundings of the Crab nebula and pulsar was calculated using the Crab pulsed fraction. The LONS intensity found is ~ 2.5 to 3 times higher than the extragalactic LONS given by Mirzoyan et al. [68], and it agrees within $\sim 7\%$ with the LONS intensity calculated using a standard reference star and a different background region (still within the camera field-of-view). This difference becomes smaller if we consider that Carramiñana et al. define the main pulse as the light curve between -0.051 to $+0.049$ phases, and the interpulse between the phases $+0.360$ to $+0.460$, so that there is a small contribution of the pulsed emission which was not considered and therefore the value of the LONS calculated will decrease, which will improve the agreement with the value measured using ζ -Tauri.

The advantage of this new method is that it relies on the stability of the clock, and is independent of other sky-techniques in the DC mode, provided that the PSF is known relative to the pixel size. The determination of a representative value of the LONS for the Crab direction is important, since the threshold energy determination, and hence the spectrum, will depend on the level of the LONS.

Chapter 4

Timing Noise and Glitches in Pulsar Observations

Arrival times of radio pulses from neutron stars are easy to record and study, and provide a wealth of information about the nature of the pulsed radio sources. Therefore it gives accurate information about the source position, pulsar parameters, the propagation of pulses through the interstellar medium, etc. Arrival time measurements can be also used to test general relativity and gravitational radiation. Irregularities in arrival times, related to the pulsar rotational rate, also give clues about the pulsar interior [2, Anderson & Itoh, 1995].

The detection of pulsed γ -rays can be done with MAGIC within a relatively short time, as discussed in Chapter 2, either for canonical and millisecond pulsars. The sensitivity is however not sufficient to enable a detection within a few seconds as reached by radio surveys. A relatively strong signal allows a period search given a few hours' data, but if the signal is marginal, the correct period parameters must be known a priori. The residual phases from the epoch are then subjected to a test for uniformity, from which the significance of a signal can be assessed.

In general, pulsars are very stable rotating stars, especially millisecond pulsars. The rotation frequency decreases steadily while the star is losing energy

and it is possible to predict the period in advance for millisecond pulsars. However, for young pulsars, the arrival times can suffer unpredictable changes that must be taken into account. These time irregularities are classified in two types: timing noise and glitches. The first one can be fitted by a n -degree polynomial in frequency and the frequency derivatives, since the irregularities are smooth, whereas glitches, the second type, are characterized by spectacular jumps in the rotational frequency as well as its derivatives.

If contemporary radio parameters are not available, we have to extrapolate archival radio parameters to the observation time in question. Such an extrapolation must be accurate enough to avoid significant pulse smearing.

In this chapter we will investigate the irregularities in the pulsar arrival times and the effect of this pulsar timing noise on γ -ray observation when attempting to extrapolate ephemerides from archival radio observations to the epoch of γ -pulsar observations.

4.1 The Basic Transformations

Before one does any folding or applying of statistical methods to the arrival time, it is necessary to transform the arrival times to an inertial frame. There are several factors which modify those arrival times: pulsar proper motion, rotation of Earth, the Earth's orbital motion around the Sun, and relativistic corrections. One needs to refer the arrival times to a steady point of reference if any calculation has to be done. Those factors can be due to intrinsic modification in the neutron star, or due to relative movements between the observation site and the pulsar:

- To transform the arrival times to an inertial frame of reference, one should transform them to the center of inertia of the Solar System, known as the barycentre. The Solar barycenter moves essentially uniformly through space and is therefore a convenient inertial frame of reference.

- The pulsar period increases steadily during the observations, so the first derivative of period P , i.e. $\dot{P}$ must be included if the observation time is long enough.
- Rotation of the Earth introduces a variable time delay of up to the transit time over one Earth radius.
- The Earth's orbit around the Sun is elliptical, not circular. Figure 4.1 shows the annual variation due to the Earth's motion around the Sun in an elliptical orbit. The projected delay times in the ecliptic plane therefore change on an annual basis.
- The gravitational potential at the Earth differs from the potential at large distance from the Sun; furthermore it varies annually through the ecliptic plane. General relativity therefore predicts a small annual variation of an Earth-bound clock rate, as compared with a reference clock in a circular orbit.
- The second-order Doppler effect, predicted by special relativity, varies as the square of the Earth's velocity: the variation of this effect due to the ellipticity of the orbit is in practice indistinguishable from the general relativistic effect in the correction above.

For most single pulsars, transformation to the barycentre of the solar system is sufficient, and the second order corrections can be ignored. The total correction of the UTC arrival times to the inertial frame can be expressed as the sum of the above mentioned corrections:

$$t_b = t + \frac{\vec{r} \cdot \hat{n}}{c} + \frac{(\vec{r} \cdot \hat{n})^2 - |r|^2}{2cd} - \frac{D}{\nu^2} + \Delta_{E\odot} + \Delta_{S\odot} \quad (4.1)$$

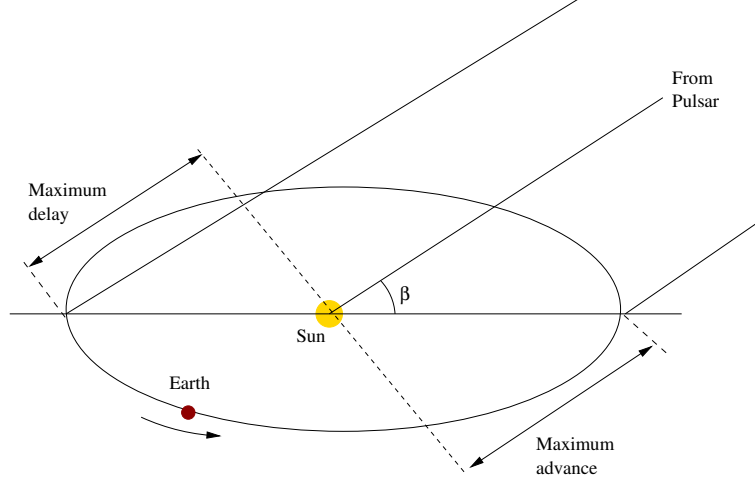


Figure 4.1: *The annual variation of pulse arrival time, assuming pulses from a pulsar at an ecliptic latitude of β .*

where t_b is the corrected arrival times at the solar system barycentre, $\vec{r}$ a vector from the barycentre to the center of the telescope, $\hat{n}$ is a unit vector pointing from the barycentre to the pulsar, c is the speed of light, d the distance to the pulsar, D the interstellar dispersion constant, ν is the rotational frequency and $\Delta_{E\odot}$ and $\Delta_{S\odot}$ are delays (Einstein delay and Shapiro delay) due to the gravitational red shift and the curvature of space time near the Sun.

Once the above transformation is performed, we can fold the arrival times with the pulsar ephemeris to determine the rotational phase $\phi(T)$. This folding process can be approximated by a Taylor expansion:

$$\phi(t) = \phi(T_o) + \nu(t - T_o) + \frac{1}{2!}\dot{\nu}(t - T_o)^2 + \frac{1}{3!}\ddot{\nu}(t - T_o)^3 + \frac{1}{4!}\ddot{\nu}(t - T_o)^4 + \dots \quad (4.2)$$

where $\phi(T_o)$ is the pulsar phase at time T_o , with ν , $\dot{\nu}$, $\ddot{\nu}$ etc the pulsar frequency and their derivatives.

We fit the arrival times with this Taylor expansion, where the irregularities are included when adding the frequency derivatives. In the next section these timing irregularities are further investigated.

4.2 Timing noise and Glitches

All known pulsars are characterised by an intrinsically precise period, modulated only by a slow monotonic increase, due to a gradual loss of rotational energy. This slowdown differs from pulsar to pulsar, depending on the pulsar activity. For instance, the change of the period of the Crab pulsar can be detected within a few hours, and the change for the Vela pulsar within days, but generally the rate of change is so small that it can only be determined from observations over a period of one year.

The rotation frequency of an isolated pulsar is expected to be decreasing steadily with time as angular momentum is carried away from the neutron star via electromagnetic radiation and particle acceleration. This spin-down torque is well-approximated by a simple power-law relation of the form:

$$I\dot{\nu} \sim -K\nu^n \quad (4.3)$$

where K is a positive constant, which depends on the magnetic moment, I is the moment of inertia of the neutron star and n is the braking index associated with the electromagnetic torque which is equal to 3 for magnetic dipole radiation, which represents the simplest pulsar model in a vacuum.

Observationally significant values have been found for a few pulsars which have been observed during a few years. For the Crab Pulsar a braking index of $n = 2.515 \pm 0.005$ was found, [64, Lyne et al, 1988]; for PSR B1509-58 the value was very similar, $n=2.8 \pm 0.2$, [54, Kaspi et al, 1994] and for PSR B0540-

69 a braking index of $n=2.01\pm0.02$, [10, Boyd et al., 1995], was determined, which is slightly smaller than the other two. The difficulty associated with such a measurement is that long observation times (without glitches) are needed to determine the braking index with high accuracy.

This slowdown can be predicted and is not a problem when we attempt to extrapolate an ephemeris from, for example, one month earlier.

Apart from this steady increase in the rotational period, some very interesting irregularities in pulsar rotation have been observed: These irregularities are believed to be related to changes in the interior of the neutron star. Therefore timing observations are a major source of information on the interior structure and long-term rotational behavior of neutron stars.

Two categories of timing irregularities exist: a general 'noisy' and fairly continuously erratic behavior, the so-called timing noise, and a more spectacular step change in rotation speed, popularly known as 'Glitch'.

Timing noise, defined as departures of the pulse arrival times from the predictions of a low-order polynomial model (Eq. 4.2), has been observed mainly in young pulsars. The level of timing noise is characterized by a deviation or r.m.s. level over a specified period of observation. This is determined, from a fit to the frequency and its first derivative, usually from a three years period of observations. The r.m.s. deviation with respect to the best fit marks the level of activity or timing noise in the pulsar. Timing noise has been attributed to a number of processes, ranging from pure random walk, micro-jumps, precession and binary orbital motions, to mixtures of two or more of these processes.

Unlike glitches, timing noise appears to be widespread, and it seems to change dramatically from pulsar to pulsar, being more active in young pulsars. It will be shown later, that very fast rotating pulsars are very stable and the timing noise does not influence the evolution of the rotational frequency significantly. Fig. 4.2 shows pulse arrival time residuals (following fits to a second order polynomial), plotted for four different pulsars in the South African radio telescope HartRAO

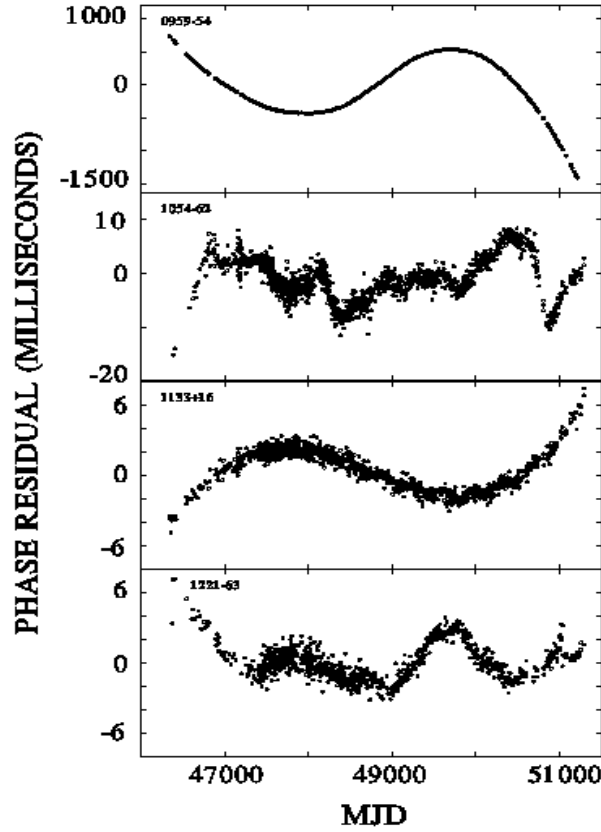


Figure 4.2: *Timing noise in 4 pulsars from observation at HartRAO telescope in South Africa during more that 10 years. The phase residual in millisecond vs. the MJD is plotted after the best fitted of arrival times.*

sample (Hartebeesthoek Radio Astronomy Observatory) [17, Chukwude et al. 2002]. The residuals in the figure represent PSR B0959-54, PSR B1054-62, PSR B1133+16 and PSR B1221-63. The phase residual, derived from observations spanning more than 10 years, differ from pulsar to pulsar although it is clear that the dominant features are represented by cubic terms. The four pulsars are relatively slow rotators with periods between 0.2 and 1.4 sec. The rotational frequency ν and frequency derivative $\dot{\nu}$ are extracted by fitting equation 4.2, with $\ddot{\nu} = 0$, in short independent segments of arrival times.

Glitches, on the other hand, are characterized by sudden increases in rotation rate, and they are, again, a common feature in young pulsars. These spin-up events are superposed on the long term rotational slowdown, and are usually followed by a relaxation towards the extrapolated pre-glitch frequency over a period of days to weeks. High glitch activity has been observed in pulsars like the Crab [104, Wong et al, 2001] and Vela pulsar, the last one with at least twelve glitches observed at radio frequencies between 1969 and 1996. Twelve glitches have also been detected from PSR J1341-6220 during just 8.2 years, making it the most frequently glitching pulsar known [103, Wang et al., 2000].

A typical glitch involves an increase in the rotational frequency of magnitude $\frac{\Delta\nu}{\nu} \sim 10^{-9} - 10^{-6}$, accompanied by an increase in $|\dot{\nu}|$, and $\frac{\Delta\dot{\nu}}{\dot{\nu}}$ of a few percent.

If the pulsar we are analysing shows glitches, it is not possible to fit the residual throughout the entire period of observation, since the ephemeris is not constant. It is then without prediction power. A new ephemeris must be derived following each glitch.

The purpose of this study is to investigate the activity or timing irregularities associated with a population of pulsars, to assess the feasibility of extrapolating radio pulsar ephemerides for online γ -ray observations. Therefore, for very active pulsars, it would be difficult to check on-line for any periodicity, unless contemporary radio observation are available.

4.3 Effect of Timing Noise and Glitches on γ -ray Pulsar Observations

The correct period parameters must be known, but if contemporary radio parameters are not available for online observations, we have to extrapolate archival radio parameters to the observation time in question. Furthermore, such an extrapolation must be accurate enough to avoid significant pulse smearing. As explained

above, timing noise and glitches play an important role, since they modify the steady slowdown in an unpredictable way.

In this section we investigate the smear in the γ -ray pulse profile during a single night of observation, given the absence of contemporary radio pulsar observations during on-line analysis when *quick look* procedures must be employed to determine the status of a candidate.

The pulsar ephemerides from the archival data of HartRAO and Princeton [65, Manchester et al, 2001] (between 1989 and 1998) provide an excellent opportunity to study the accuracy of extrapolations of such ephemerides to the present moment, if an appropriate time shift is introduced.

For our studies, the HartRAO data base provided ephemerides for 5 pulsars and 43 were selected from Princeton data base. The long duration of these data bases allows us to study the effect of timing noise when attempting to extrapolate ephemerides in the absence of contemporary radio observations.

4.4 Testing the extrapolation of ephemerides

MAGIC was inaugurated on the 9th of September 2003 and after the commissioning period it will be ready for γ -ray pulsars observations. Since the last year with pulsar observation given in the present cataloguers is 1997, the question remains if it is possible to extrapolate radio parameters derived from around 1997 to an observation period during 2003.

The smear in the phase due to an inaccurate frequency, can be quantified through an independent Fourier spacing (IFS). Therefore, when evaluating periodicity searches, the concept of an independent Fourier spacing plays an important role [28, de Jager, 1987]. It can be defined as follow:

If T_{obs} is the total length of a data sample, which is usually $T_{obs} = 6$ hours for a typical single night of IACT γ -ray observations, and ν_1 and ν_2 represents two

different frequencies, then δ , defined by the expression

$$\nu_2 - \nu_1 = \frac{\delta}{T_{obs}}, \quad (4.4)$$

will represent a measure of independency between ν_1 and ν_2 . The value of $|\delta| = 1$ corresponds to one independent Fourier spacing and the set of frequencies spaced by $1/T_{obs}$ from each other are independent.

The uncertainty or resolution in the period of any significant radiation found at a specific period is usually 1 IFS. To investigate the phase smear, we search for periodicity in intervals of 1 IFS.

If we neglect any frequency second derivative, the rotational phase from Eq. 4.2 can be expressed as:

$$\phi = \nu(t_i - t_0) + \frac{1}{2}\dot{\nu}(t_i - t_0)^2, i = 1, \dots, N \quad (4.5)$$

By taking $T_{obs} = (t_N - t_0)$ as the total duration of the observations, the phase shift $\Delta\phi$, due to a change in frequency ν alone (due to timing noise or glitches), would be:

$$\Delta\phi + \phi = (\nu + \Delta\nu) \cdot T_{obs} \quad (4.6)$$

This means, if $\Delta\phi = 1$, the phase shift from a total observation time T is one complete cycle, so that the signal would be completely smeared in phase,

Rewriting Eq. 4.5 in terms of δ gives essentially the same expression as Eq. 4.4

$$\delta = \Delta\nu \cdot T_{obs} \quad (4.7)$$

If $\Delta\nu = \nu_{extrap} - \nu$, with ν_{extrap} the extrapolated frequency, one can calculate the δ value and so evaluate the validity of the extrapolation.

Since most of the entries in the Princeton catalogue terminate around the year 1997, and assuming that MAGIC observations start during 2003 - the inauguration year of the MAGIC telescope, the feasibility of extrapolating parameters from 1997 to 2003 can be assessed by studying the accuracy of predicting 1997 data from data taken ~ 6 years earlier. This means that data from 1990 to 1992 may be used for this purpose. The extrapolated frequency is given by:

$$\nu_{extrap} = \nu_o + \dot{\nu}(t - t_o) + \frac{1}{2}\ddot{\nu}(t - t_o)^2, \quad (4.8)$$

where t_o is the reference time and t the time corresponding to the observations in question. The difference between the true frequency $\nu(t)$ at the present epoch t and the extrapolated frequency defines the phase smear for the γ -ray pulse in terms of Eq. 4.7, with $\Delta = \nu - \nu_{extrap}$.

Given the large combined data base of timing data provided by Princeton and HartRAO, we can make a statistical study of the accuracy of extrapolating timing solutions in general. The effect of glitches will be implicitly included in these studies since we are studying the evolution of the rotational period over an interval of several years.

Fig.4.3 shows the phase smear δ for two of the most active pulsars known: Crab and Vela over a 5-year period, while Fig.4.4 illustrates the evolution in time (MJD) for a quiet pulsar such as PSR B0136+57. For both Crab and Vela the extrapolation is not reliable, because of the large timing noise and the frequent glitches, while for the third pulsar it is clear that we do not need contemporary radio observation since the period is stable enough for a 6 hours observation time during a few years.

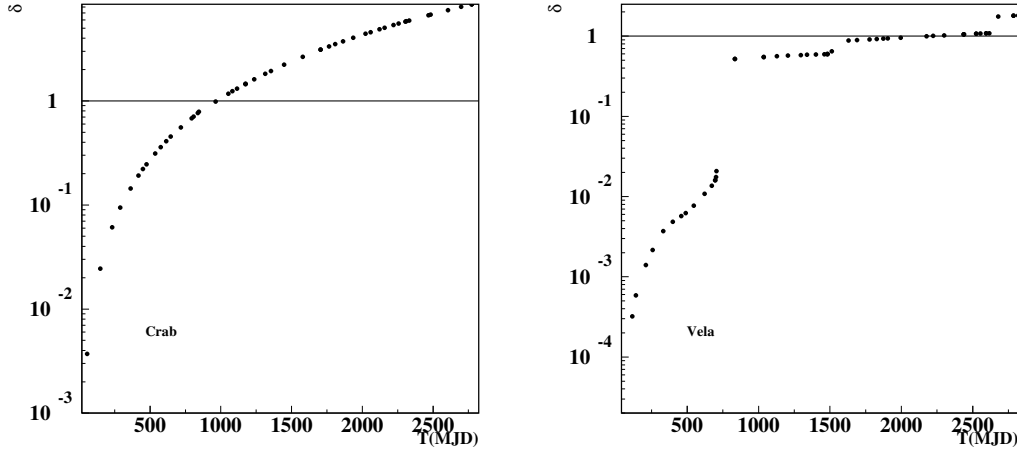


Figure 4.3: *Phase smear δ vs. time $T = (t_i - t_o)$ in MJD for two pulsars with significant noise, for a observation time $T_{obs} = 6$ h: the Crab pulsar on the left, and the Vela pulsar on the right. Extrapolation over a period of more than two years lead to a phase smear of more than 10%.*

To study the extrapolation accuracy versus pulsar parameters such as period and spindown power $\dot{E}$, the phase smear (Eq. 4.7) was calculated for $T = 10^3$ days, which represents the minimal available time span of radio observations. Fig. 4.5 shows the phase smear in terms of $\log(\delta)$ versus $\log(\text{Period})$, whereas Fig. 4.6 shows the same against $\log(\dot{E})$. The horizontal dashed lines represent 100% phase smearing ($\log\delta = 0$), 10% phase smearing as allowed for broad pulse profile ($\log\delta = -1$) and 1% smearing ($\log\delta = -2$), representing the maximal phase smear allowed for a sharp Crab & Vela like pulsars.

It is clear that short period, high $\dot{E}$ canonical pulsars (excluding millisecond pulsars) result in maximal phase smear, whereas the older, lower $\dot{E}$ pulsars reveal a quieter behavior.

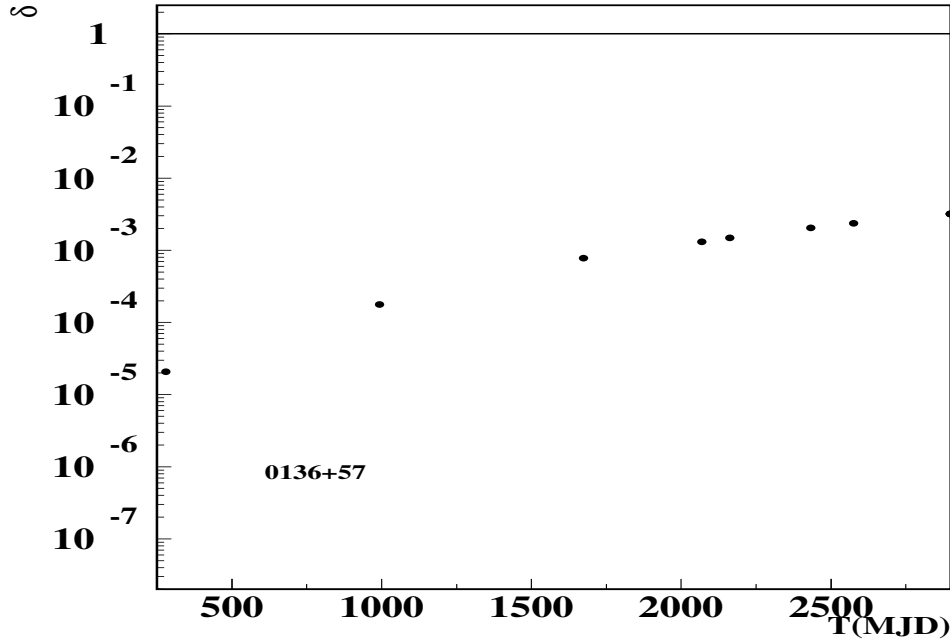


Figure 4.4: *Number of IFS vs. MJD for a quiet pulsar: PSR B0136+57. Even for the largest extrapolation, which is approximately 6.8 years, $\delta = 10^{-3}$. That means we will find the periodicity using ephemeris from even 7 years ago.*

It is also clear that observation of Crab & Vela require most regular updates for their frequency parameters, but with the central pixel available for Crab, accurate timing information from such optical observation will always allow an unambiguous search for γ -ray pulsations.

It is also important to know the maximum integration time allowable, until the pulse will be smeared. Fig.4.7 shows the number of hours needed, allowing a smear in the phase of 1% (dots) and of 10% (crosses). The four horizontal lines mark 6, 48 hours, 2 weeks (one observation shift) and 3 month (observation season) respectively. The number of hours is calculated from the expression 4.7 assuming an extrapolation period of $(t - t_o) = 10^3$ days. It is remarkable that for some pulsars with low spin-down power, the stability of the period is sufficient to allow the folding with extrapolated ephemerides for $T_{obs} = 10^6$ hours.

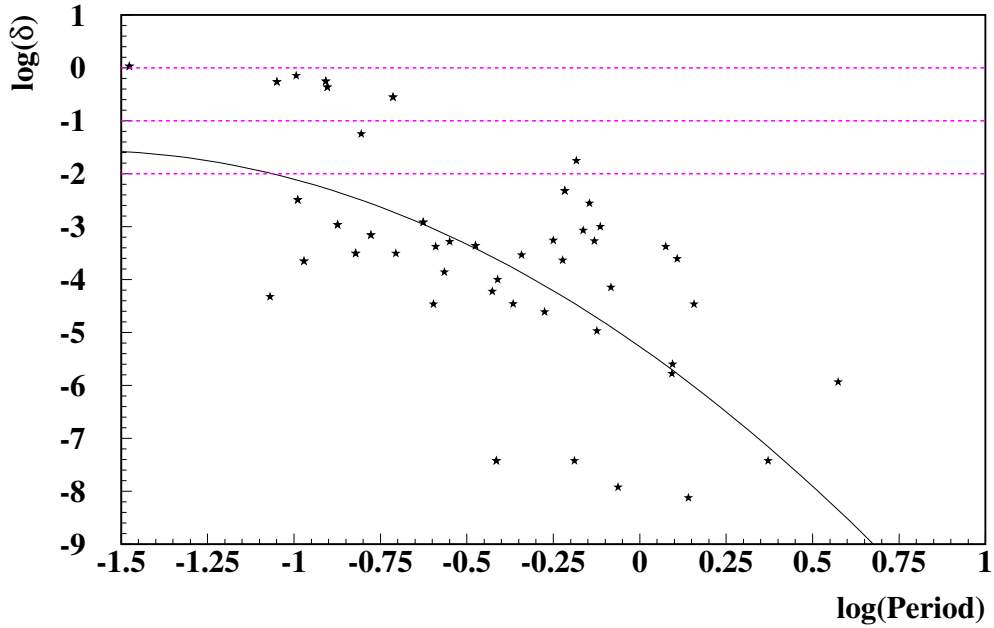


Figure 4.5: $\log$ of δ values in units of IFS vs. the $\log$ of the period for a extrapolation time of 10^3 days in seconds and 6 hours observation time. The Crab pulsar shows up in the upper left corner.

4.5 Conclusions

As a conclusion of these studies, it is clear that it is possible to extrapolate the radio parameters of most radio pulsars over several years.

For a given period or spin-down power we can assess from statistical arguments the minimum observation time T_{obs} for which the extrapolation of the ephemerides will be valid.

From 48 pulsars analysed and allowing a phase smear of 1%, for an observation time of 6 hours, we would be able to extrapolate the ephemeris of 83% of 48 pulsars from radio data taken ~ 3 years ago. For $T_{obs} = 48$ hours it would be possible to use the extrapolations for 77%, and for $T = 2$ weeks, 42% of the

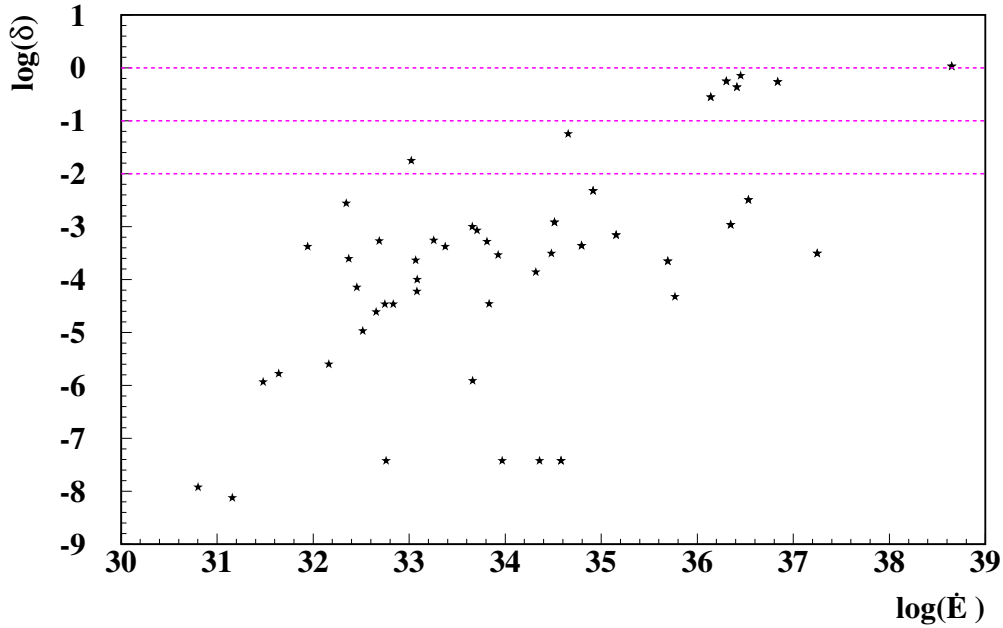


Figure 4.6: *log of δ values in units of IFS vs. the log of the spin down energy. The horizontal lines mark limits to allow the extrapolation for different pulsar duty cycle. Again the extrapolation time is 10^3 and $T_{obs} = 6$ hours.*

total number of studied pulsars could be analysed without contemporary radio observations. If we consider 3 months, then only for 25% pulsars the extrapolation is possible.

For online observation with MAGIC, for pulsars with periods longer than ~ 0.3 s, it is possible to do a search from a 3 years old catalogue, and use the extrapolated frequency without loosing the periodic signal. The same can be said about millisecond pulsars, although, for fast rotating as well as young pulsars, simultaneous observations are required. In some particular cases, when the pulsar is observable at optical wavelengths, the camera central pixel can be used to determine the periodicity, as demonstrated in Chapter 3. Unfortunately (see Chapter 3) only 6 pulsars so far show optical pulsed emission. The Crab pulsar is one of them and also one of the most active pulsars in terms of timing noise, but its rotational frequency and pulse phase remain constant throughout

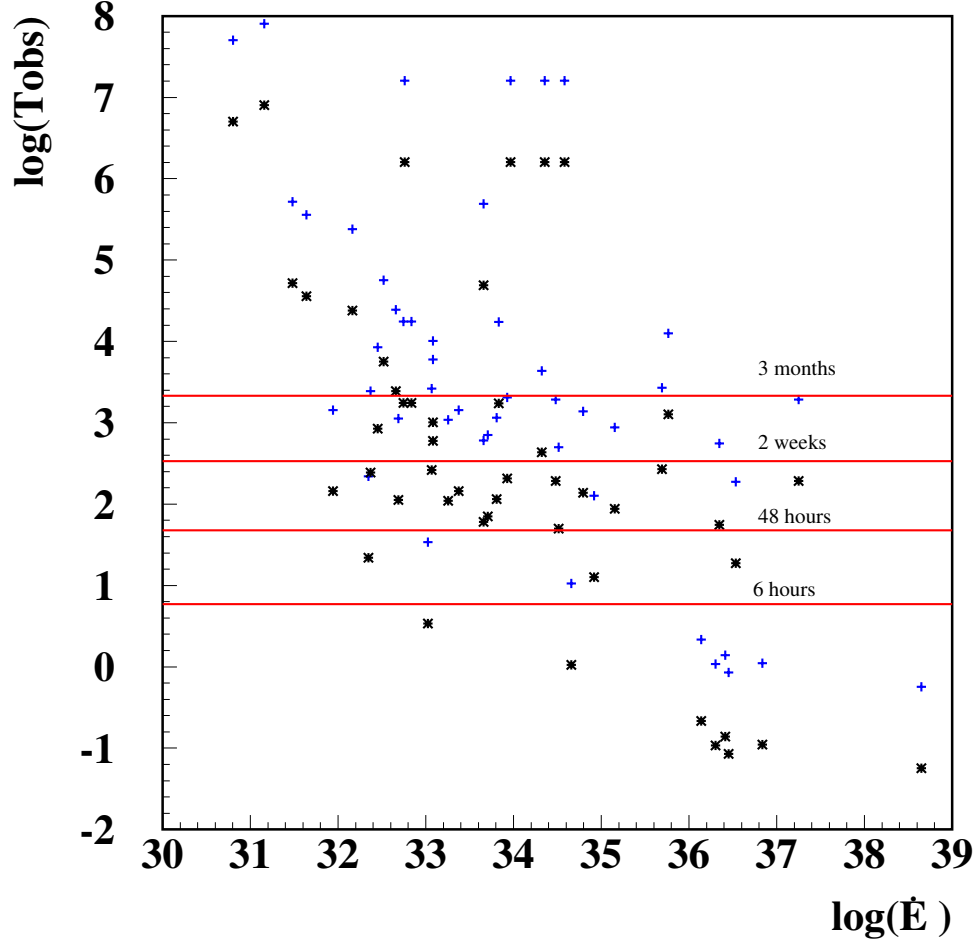


Figure 4.7: *The log of the maximum integration time in hours, allowing a smear in the phase of 1% (dots) and 10% (crosses) vs. the spin-down power. The horizontal red line on the top marks the limit from the top to the bottom: for 3 months (observation season), 2 weeks (observation shift), 48 hours, and 6 hours.*

all wavelengths. So thus contemporary optical observation can be used for γ -ray observations.

Chapter 5

A 2 to 10 GeV γ -rays detector

5.1 Introduction

The EGRET detector on board the COMPTON Gamma-Ray Observatory [95, Thompson et al., 1993] could reject the hadronic background efficiently using its anti-coincidence shield, but was rate-limited as a result of its limited collection area of $\sim 0.1 \text{ m}^2$ up to about 30 GeV. Since most GeV sources are weak, with threshold fluxes around a few times $10^{-4} \text{ m}^{-2} \text{ s}^{-1}$ above 1 GeV [58, Lamb & Macomb, 2000], typical space platforms, such as GLAST with a collection area of $\sim 1 \text{ m}^2$, have to integrate for at least 10^4 s to capture a few photons above a GeV from a region free of diffuse γ -ray emission.

Above 10 GeV, typical detection times with GLAST increase to at least 24 hours. For steady sources this constraint is offset by the large sky coverage of this detector [35, Dingus et al., 1997], but fast transient phenomena at some GeV call for much larger collection areas.

There is sufficient physics motivation for observations at an energy of a few GeV: most canonical pulsars show spectral cutoffs in this range, whereas General Relativistic Electrodynamics predict hard spectra up to $\sim 100 \text{ GeV}$ from millisecond pulsars, as discussed in Chapter 2. Large redshift AGN observations are

also not hampered by intergalactic absorption for photon energies of a few GeV [90, Salamon & Stecker]. Finally, the GeV range is also important for Particle Astrophysics [40, Flix et al. 2003], [8, Blanch et al., 2003], [61, López et al., 2003].

To achieve the aforementioned goals with ground-based telescopes, an array of MAGIC-type telescopes at an altitude of 5 km was suggested by Aharonian et al. (2001) with the purpose of achieving a large collection area at 5 GeV with good background rejection capability. An alternative design is the so-called ECO-1000 telescope [67, Merck et al., 2003], one single dish with an area of 1000 m², at the MAGIC site, that is, a moderate altitude of 2200 m. These instruments would provide excellent spectral response from a few GeV to multiple TeV energies, if the normal Imaging Atmospheric Cerenkov Technique [100, Weekes, 1989] can be applied. However they require significant capital investment. The array also necessitates special infrastructure adapted to high altitudes.

For the immediate future we offer an alternative detector with collection area up to 2000 m² at 5 GeV, using the existing 17-meter class MAGIC Telescope and its present $\sim 25 - 30\%$ quantum efficiency photomultipliers - [20, Cortina et al., 2003]). The rationale, advantages and limitations are the following:

(a) the angular resolution is limited to ≥ 0.36 degree, with the latter radius defined by a cluster of 7 hexagonal pixels in the center of the camera. This cluster defines a Cerenkov trigger, but consisting of those pixels without bright stellar content, that is, excluding the brightest pixels due to stellar light.

(b) we show that the mean number of photoelectrons from a 5 GeV on-axis γ -ray shower is slightly larger than the mean number of the night sky background (LONS) photoelectrons if a 0.36 degree trigger radius at the center of the camera is selected (this statement holds for an average LONS level of $(1.7 \pm 0.4) \times 10^{12} ph \cdot m^{-2} \cdot s^{-1} \cdot sr^{-1}$). This is expected to hold for any cluster for which the normal direction to the cluster corresponds to the source direction;

(c) for such a centralized cluster, we find that the RMS of the number of

photoelectrons from a 5 GeV γ -ray shower is comparable to the mean number of photoelectrons per shower, whereas the corresponding RMS for the LONS is always equal to the square root of the mean. Furthermore, since the mean photoelectron level from such low energy showers is comparable to, or slightly larger than the mean LONS level, we can set a trigger threshold well above the mean LONS level, such that the accidental trigger rate on LONS events is negligible, whereas an appreciable trigger rate of 5 GeV showers can still be achieved, as allowed by the large fluctuations around the mean ph.e. level.

(d) the main background contribution arises from proton triggers inside the same 0.36 degree trigger diameter, which is about 60 Hz if no size cut is accepted, but decreases to ~ 6 Hz if we restrict to γ -ray energy smaller than 10 GeV. This limits this special “pulsar trigger mode” to EGRET energies. This energy range complements the region from 10 to 100 GeV which is covered by IACT. These IACTs are equipped with 4 to 5 degree FOV pixelized imaging cameras that were designed for image analysis as suggested originally by Weekes et al. (1989). The proton background rate of 6 Hz can be further reduced by using the rest of the camera in anti-coincidence trigger.

(e) the collection area around 5 GeV is limited by the total LONS in the abovementioned cluster. Whereas a collection area of $\sim 1000 \text{ m}^2$ may be reached at 5 GeV for the present MAGIC Telescope, a further factor 10 increase is achieved if the LONS is reduced by a factor of 3 in the central 6 pixels. This demonstrates the necessity of removing at least the brightest pixel from the cluster in a dynamic feedback loop as discussed previously. From a physics viewpoint, we expect extragalactic sources to yield the best sensitivity, given the lowest LONS level above the galactic plane.

The performance presented below is based on showers generated by CORSIKA [53, Knapp et al., 1998].

5.2 Proposed Technique

The MAGIC camera consists in 577 photomultipliers (PMs) in a hexagonal frame (see Fig. 5.1). Each PM has a quantum efficiency of $QE \sim 25 - 30\%$. The mean diameter of each of the central 397 pixels is 0.11° . The central pixel (numbered "0") is reserved for optical studies as discussed in see Chapter 3, but will also be employed for γ -ray triggers. Such a HPD, which serves as a single photon counter can also serve as a fast Cerenkov detector, as well as an optical photometer.

The high LONS background in the whole camera, combined with the constraints in using the imaging technique for γ /hadron discrimination at low energies, makes it difficult to reduce the energy threshold to very low energies for this type of Cerenkov telescopes. In fact the actual threshold is ~ 30 GeV in MAGIC phase I and ~ 10 GeV in phase II [66, Martinez, 2003] (with a camera of high quantum efficiency PMs).

A novel technique is proposed in this Chapter using only the center of the camera to detect very low energy γ -ray showers, that is, γ -ray showers with energies less than 10 GeV.

For that purpose we restrict our interest to the light detectors in the center of the camera, to be more precise to the central pixel (PM #0) and the so-called central ring: the pixels around the central pixel (PM # i , $i=1,\dots,6$ in the standard camera numeration) as shown in Fig. 5.1.

This figure shows a schematic view of the camera and the central ring pixels. As already mentioned each PM has an average diameter of 0.11° . The diameter of the Cerenkov light detector is therefore restricted to 0.36° to be compared to the 3.5° diameter of the total camera.

This narrow field of view allows us to reject most of the isotropic background. Moreover, showers at low energies hit the camera mainly in this area since they develop higher in the atmosphere. The summed signal from six of the seven pixels (the brightest of the seven pixels is rejected) will be used to determine the

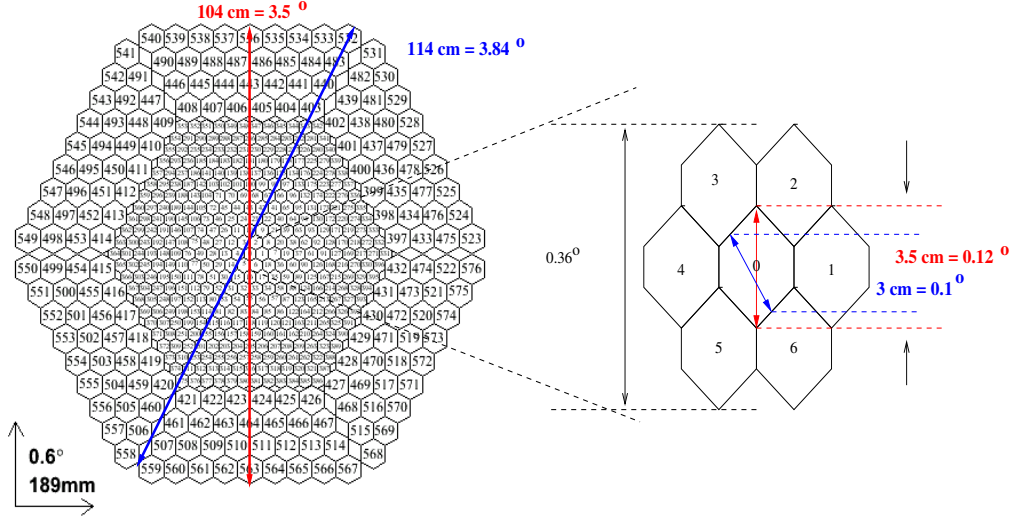


Figure 5.1: *Schematic view of the MAGIC camera with 577 pixels. The inner camera consists of 397 pixels with 0.11° diameter, while in the outer part, the pixel size is $\sim 0.22^\circ$. On the right we zoom into the center of the camera (maximum diameter 0.36° .)*

sensitivity for 2 to 10 GeV sources. For this particular study, the central pixel will be excluded. This will account for γ -ray sources which are optically bright.

The central pixel and the central ring are working independently at the moment but we expect to be able to combine them when observing γ -ray candidates without optical counterparts. The central pixel is equipped with an HPD of high quantum efficiency ($\sim 43\%$). The signal will be amplified and split in two branches: The *fast Cerenkov line*, which will be used for 2-10 GeV γ -ray detection (P_o in figure 5.2a), and the slow line for optical monitoring. A discriminator will be set up with a given threshold to register single photon counts.

For the purpose of this study we will assume that the QE of the central ring (as for the rest of the camera) is lower ($\sim 30\%$) than that of the HPD. In the next future the six central ring PMs could be replaced with HPDs (E. Lorenz, 2003, private communication).

The signal received from the six PMs that constitute the central ring (with 30% QE) is amplified and then divided into two branches: one for *standard* γ -ray

observations (from 30 GeV to several TeV), named N_i , which represents all the variables stored for a pixel i , with $i = 1, \dots, 6$, (e.g. charge, time...) and, another line for the low energy γ -ray output (see Fig. 5.2b). The analog signal from each pixel P_i is summed with no time delay to give a total intensity P_{tot} . The width of the gate was defined as $\tau = 3 \text{ ns}$. The trigger threshold can be set using a discriminator. The discriminator will be set at a level that makes the accidental trigger rate induced by the LONS negligible.

5.3 Simulation studies

Simulations were performed to investigate the feasibility of the technique. CORSIKA [53, Knapp et al., 1998] was used for the Monte Carlo simulation of low energy γ -ray showers, and the proton background. The output of the camera was analysed using the MARS software, modified to count FADC counts in the central ring pixels. The relation between ph.e and FADC counts was checked and it agrees with the value given by Bigongiari et al. [7, Bigongiari et al., 2003] (1 ph.e.=3.3 ADC counts).

About $\sim 1 \times 10^4$ (MAGIC group in ETHZ, Zurich) γ -ray showers were simulated with energies between 2 and 10 GeV. The simulations presented in this section were performed for zenith and azimuth angle equal to 0° , except for section 5.5, where a few showers were simulated for 20° zenith angle and 180° azimuth angle to test the effect of the Earth's magnetic field. Only 55% of these showers produced more than 1 ph.e in the camera. Figure 5.3 shows the distribution of the number of ph.e. for showers at 5 GeV in this data set with the first bin at 0 ph.e. representing 45% of all 5 GeV showers without any charge. The distribution average for the whole camera is ~ 5.4 , which means, scaling with the $\sim 85\%$ MAGIC reflectivity, 30% quantum efficiency of the PMs and 234 m^2 mirror area of the telescope, we find an average photons $\sim 0.27 \text{ ph} \cdot \text{m}^{-2}$ at 5 GeV. This value agrees with the value estimated previously by other authors

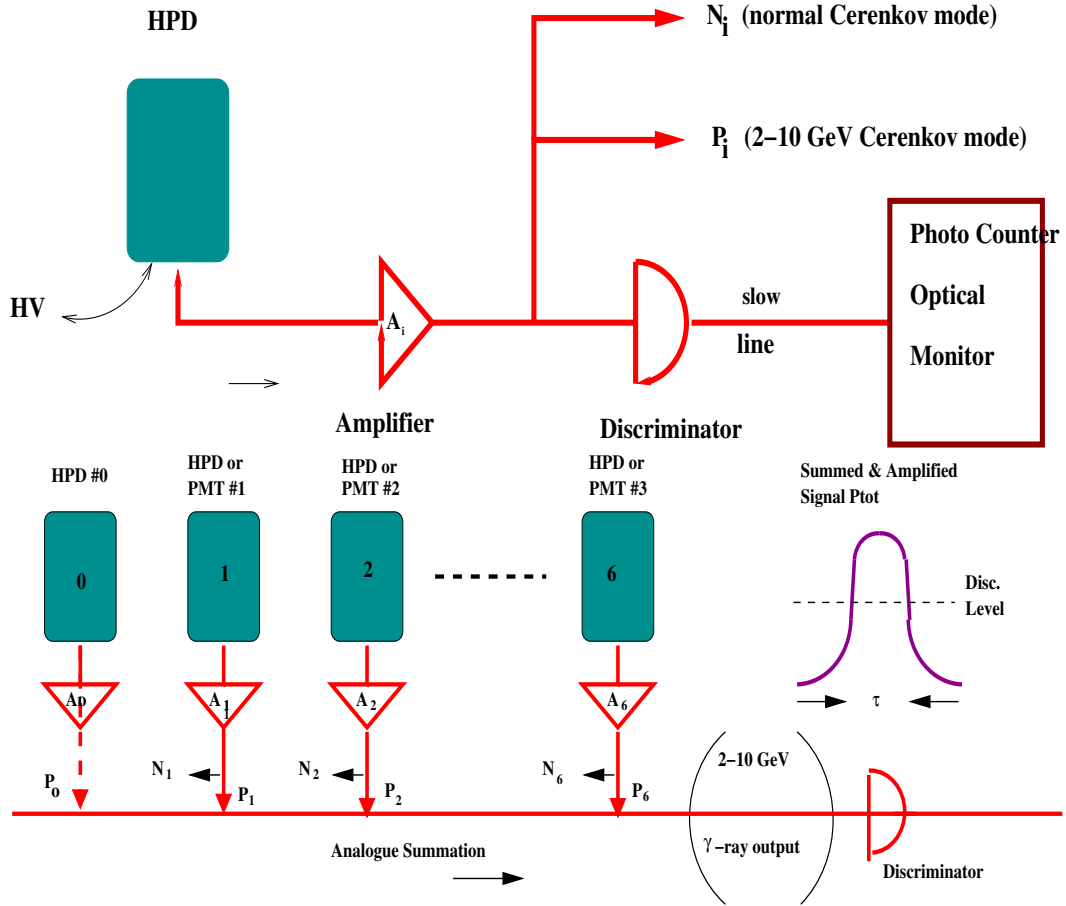


Figure 5.2: *Top: a) Configuration for the central pixel or any of the remaining central ring pixels ($i=0, \dots, 6$) represented by a HPD or at least a PM with a 30% QE (as used for these simulations). With three outputs: slow line for photon counting (optical photometry), the fast cerenkov line P_i for the 2–10 GeV mode, as well as the normal Cerenkov mode N_i , to accomodate the normal IACT presently employed by MAGIC. Bottom: b) The six pixels from the central ring P_i with a schematic view of the electronic setup. The central pixel will be included in future. The signals from the PMs will be added in analog and fed to a discriminator with a preset the trigger level.*

(see e.g. [56, Konopelko, 1997]).

The mean of the distribution of ph.e. produced in the first ring of pixels (see Fig. 5.3) is 1.7 ph.e. The RMS is 1.4 ph.e. It should be noted that the RMS of the distribution is comparable to the mean [93, Sobczynska, 2003].

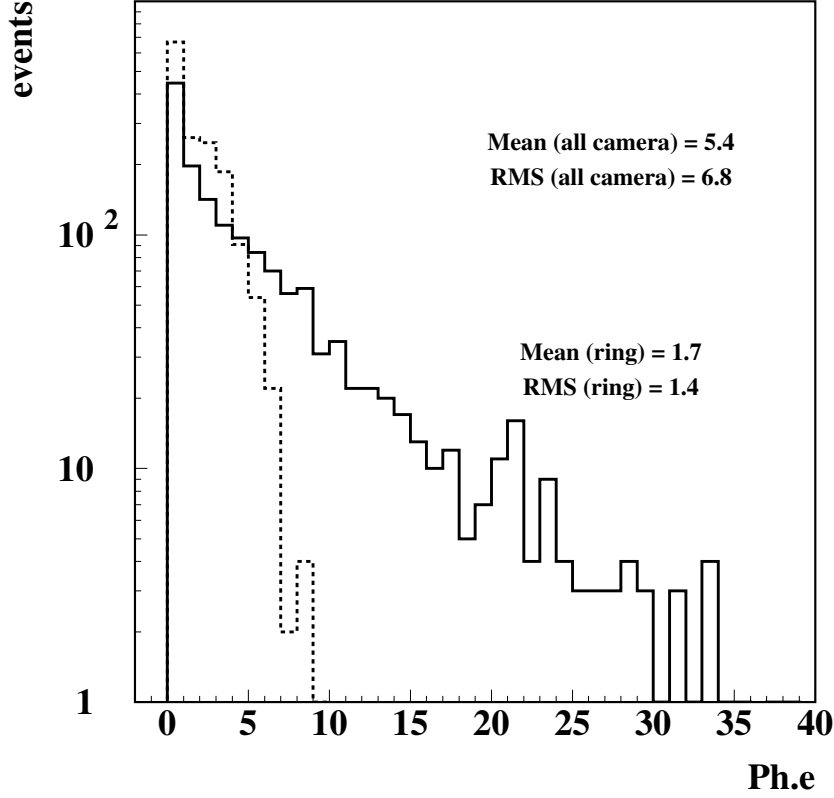


Figure 5.3: *Distribution of all ph.e. at the telescope camera for 5 GeV γ -ray showers (solid line) and distribution of ph.e. that hit the central ring (dashed line). The mean value (~ 5.4 ph.e.) corresponds to a Cerenkov photon density of 0.27 ph/m².*

Fig. 5.4 shows the ratio of the number of ph.e. in the central ring (N_{ring}) relative to the total number of ph.e. in the camera (N_{camera}) as a function of the primary energy. The ratio N_{ring}/N_{camera} increases towards lower energies. The geomagnetic field works against this trend and tends to spread the shower over the entire camera (see Section 5 for details). The behaviour observed in Fig.5.4, together with the fact that the photon density increases with increasing primary energy, is expected to produce an almost constant effective area in the 2 to 10 GeV range. The simulation below confirms this expectation.

The value of the extragalactic LONS in [68, Mirzoyan et al.,1994]; [81, Preu β

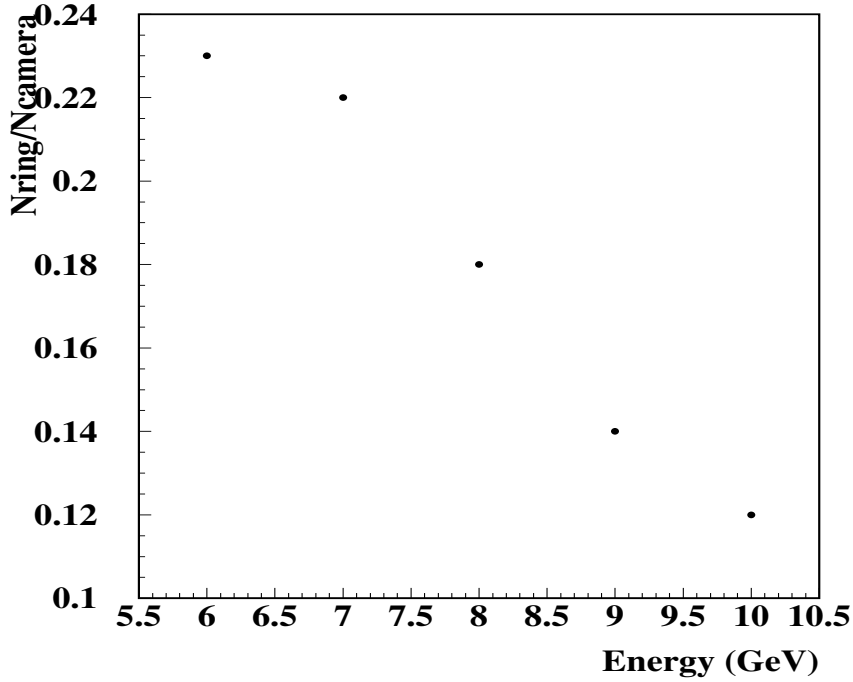


Figure 5.4: *The ratio of the number of ph.e. in the central ring and the total number of ph.e. in the camera as a function of the γ -ray primary energy. (Simulations courtesy of D. Sobczynska, 2003, for independent verification).*

et al., 2003] was included in the simulation.

10^3 γ -ray showers of 5 GeV primary energy with impact parameters selected at random in the 0 to 300 m range were generated. The total charge P_{tot} in the central ring was calculated for each shower, taking the effect of mirror reflectivity and QE into account. The distributions of the number of ph.e. induced by the LONS and the LONS plus the γ -ray showers are shown in Fig.5.5. The vertical axis shows the probability of finding a shower above a given number of ph.e. so that the probability for accidental triggers on the LONS can be properly derived.

To determine the discriminator level, we have to assume a LONS intensity. Using the value given by Mirzoyan et al for an extragalactic source:

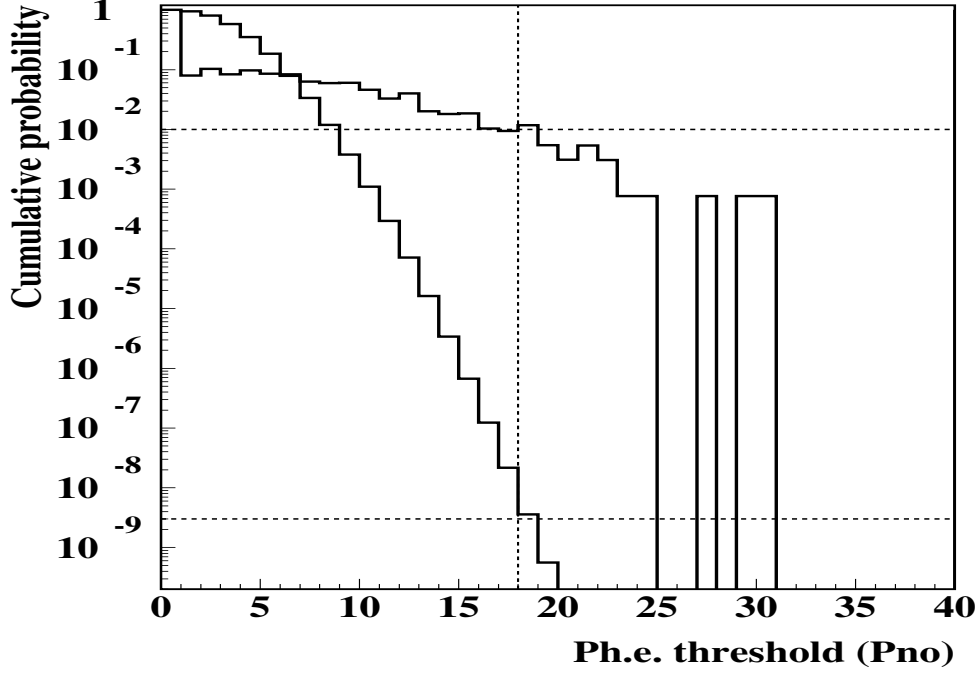


Figure 5.5: *Distribution of number of ph.e. produced by 5 GeV showers. The RMS of the signal is very similar to the mean, so the number of showers that survives after threshold $Ph.e.=18$ is still quite high. A LONS intensity corresponding to an extragalactic source is also included in the simulations. The horizontal dotted line corresponds to a LONS accidental trigger rate of 1 Hz, which is equal to a probability of 3×10^{-9} . It defines a trigger threshold of 18 ph.e., marked with a vertical dotted line. The horizontal line at this cumulative LONS of 3×10^{-9} corresponds to a trigger probability of 1% on 5 GeV showers (horizontal line on the top).*

$$I_{LONS} = 1.7 \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \text{s}^{-1} \text{sr}^{-1} \quad (5.1)$$

the LONS rate per pixel λ_{LONS} in a gate time of $\tau = 3 \text{ ns}$ will be:

$$\lambda_{LONS} = I_{LONS} \cdot A_{MAGIC} \cdot \eta \cdot QE \cdot \Omega \cdot \tau = 0.58 \text{ ph.e} \quad (5.2)$$

where A_{MAGIC} is the reflector area, η the mirror reflectivity, QE the PMs quantum efficiency and Ω the solid angle subtended by 1 pixel.

Because only six pixels are used, the average number of ph.e. in the ring is $6 \cdot \lambda_{LONS} \sim 3.5 \text{ ph.e}$ for an extragalactic LONS. The gaussian distribution of the corresponding number of ph.e. is also shown in Fig. 5.5.

We select a trigger threshold P_{no} to avoid most of the accidental triggers. The accidental trigger rate B_{LONS} can be expressed as:

$$B_{LONS} = \frac{\tau}{\text{Prob}(\geq P_{no})}, \quad (5.3)$$

and the Poisson probability:

$$\text{Prob}(\geq P_{no}) = \sum_{i=P_{no}}^{\infty} \frac{e^{-6 \cdot \lambda_{LONS}} \lambda^i}{i!}, \quad (5.4)$$

as shown in Fig. 5.5 for an average LONS ph.e. level of 3.5 ph.e. Several LONS accidental rates, 1 Hz, 0.1 Hz and 0.01 Hz can be allowed, depending on the intensity of the LONS around the γ -ray source. The arrival times and pulse height P_{tot} are recorded if the trigger condition ($P_{tot} \geq P_{no}$) is fulfilled.

The rest of ph.e. in the entire camera are used for the normal operation of the IACT [100, Weekes, 1989]. The charges stored in N_1 to N_6 in the central ring, combined with the charge in the rest of the pixels N_7 to N_{577} are used to calculate the standard Hillas parameters. This is not relevant for the discussion below, although this information can be used either as anti-coincidence or to define other similar clusters in the camera.

Applying a certain threshold, the trigger efficiency will be given by the expression 2.26 and the corresponding effective collection area can be calculated. Figure 5.6 shows the effective area as a function of the primary energy between 2 and 10 GeV for two different accidental trigger rates: 1 Hz (bullets) and 0.01 Hz

(stars). Depending on the LONS level at the source location, we will select either of these accidental trigger rates.

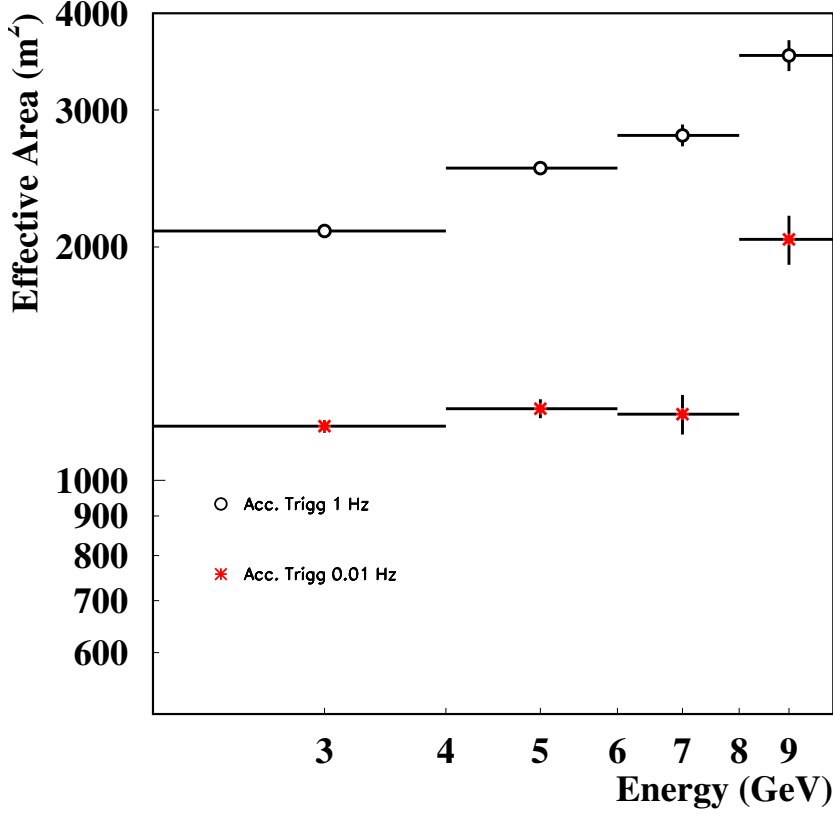


Figure 5.6: *Effective area between 2 and 10 GeV using the central ring technique. The effective area reaches 3000 m^2 for an accidental trigger rate of 1 Hz and $\sim 1500 m^2$ for an accidental trigger rate of 0.01 Hz.*

A mean value of 2000 m^2 for the effective area is reached for extragalactic γ -ray sources. Fig. 5.7 compares the effective area using our "very low energy" mode from 2 to 10 GeV with the standard operation mode [32, de los Reyes, 2002] (see Fig. 2.6). The effective area of other γ -ray experiments is also plotted in the figure. EGRET has a mean effective area in the 1 to 10 GeV range of $\sim 10^{-2} m^2$ [96, Thompson et al., 1993] and GLAST is expected to reach $\sim 1 m^2$ above 1 GeV.

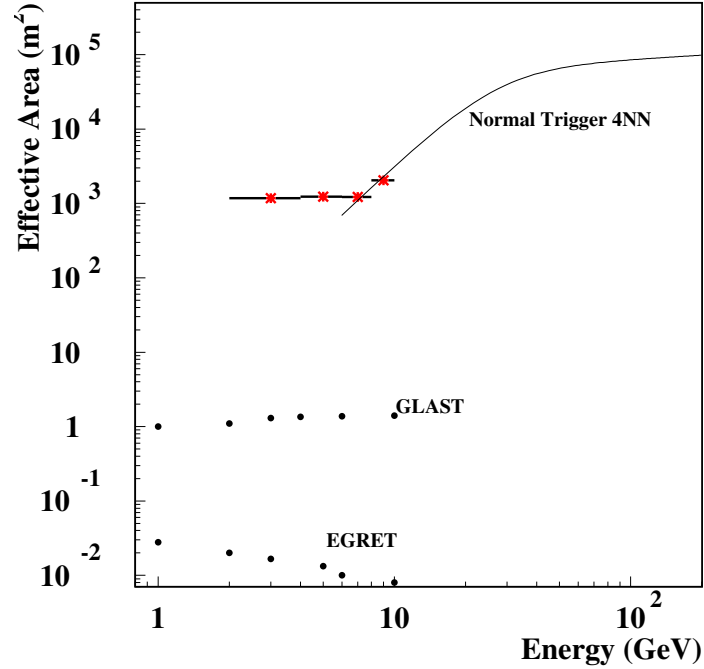


Figure 5.7: *Effective area using two different trigger conditions: central ring (star) and 4NN coincidence of pixels with more than 3.5 ph.e (solid curve indicated "Normal Trigger 4NN"). The effective areas of the two main satellite detectors working in the same range of energies, EGRET and GLAST, are also represented.*

Contrary to the standard effective collection area, the new effective area defined by the central ring technique does not decrease dramatically for low energies, and it is three orders of magnitude larger than the future GLAST satellite, although the energy resolution is very poor compared to GLAST.

The flat response of the area vs. energy in the 2 to 10 GeV range was anticipated based on the increasing compactness of the shower with decreasing energy as illustrated in Fig.5.4, combined with an increasing photon density with increasing energy.

However the effective area depends on the trigger level we require, and this

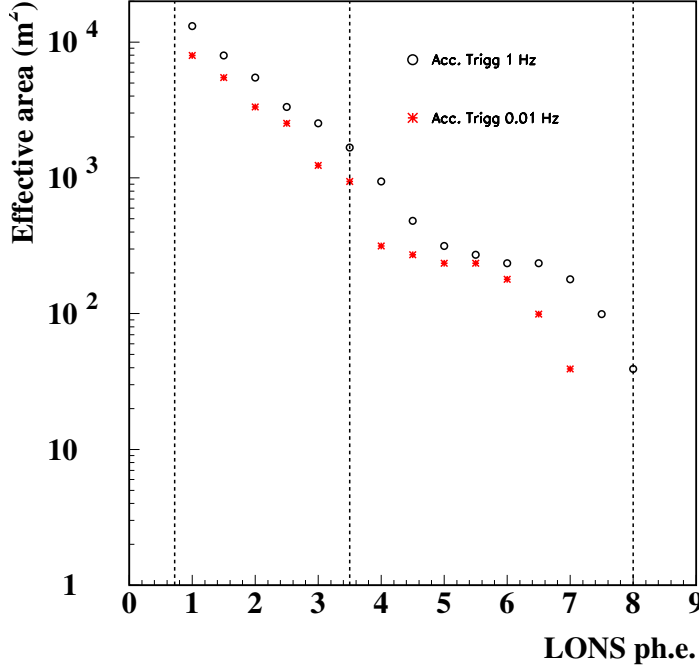


Figure 5.8: *Effective area at 5 GeV using the central ring technique vs. the night sky background level. The empty circles show the area for 0.01 Hz accidental trigger rate and the red crosses for 1 Hz accidental trigger rate. The three vertical lines mark the average number of ph.e. for 3 different LONS conditions (expressed as the sum in the ring of 6 pixels): 0.7 ph.e. for diffuse LONS in the absence of stars, 3.5 ph.e. for an extragalactic source and ~ 8 ph.e. for the LONS around the Crab pulsar, a galactic source.*

trigger level is determined by the LONS associated with the source position.

Fig. 5.8 shows the effective area at 5 GeV as a function of the average ph.e level from the LONS. The effective area is calculated allowing 0.01 Hz (one accidental trigger in 100 s) of accidental trigger, and for 1 Hz, (one accidental trigger in 1 second).

For a more realistic result, we calculate the response of the 6 PMs of the central ring for three different values of the LONS. The standard value used by the MAGIC collaboration in the simulation of the camera response corresponds to diffuse LONS with no stars and is equal to 0.12 ph.e. per pixel. For the 6 central

ring pixels it adds up to 0.72 ph.e. Using an extragalactic measure of the LONS, the number of ph.e is around 3.5 ph.e. Using the value given by Oña-Wilhelmi et al, [78, Oña-Wilhelmi et al.,2003], discussed in Chapter 3, for the LONS around the Crab pulsar, a galactic source with $I_{LONS} = 4.6 \times 10^{12} \text{ ph. } m^{-2} s^{-1} sr^{-1}$, will produce ~ 8 ph.e in the central ring. These three values of the average number of ph.e in the central ring due to the LONS, are also marked in Fig. 5.8 with vertical lines. They correspond to an effective area of $1.3 \times 10^4 \text{ m}^2$, $2 \times 10^3 \text{ m}^2$ and 50 m^2 respectively.

For galactic sources and due to the high night sky background the effective area is very small or even negligible if the LONS level is very high. However in this range of energy the γ -ray intensity from pulsars is relatively large, so collection areas as small as 50 m^2 are still useful provided that the cosmic background is low enough. Apart from the accidental triggers due to the LONS, the effect of the cosmic ray electron and proton background must also be estimated.

5.3.1 Proton Background Rejection

One of the main problems at low energies using the standard 4NN trigger condition is that the light density is small. The Hillas parameters are hence difficult to evaluate and the γ /hadron discrimination is not possible anymore using normal imaging techniques.

For the central trigger technique, the narrow field of view may reduce the background. Some estimates have been made to calculate the expected proton rate. The rejection of this background is discussed below.

Proton showers for energies between 20 and 2000 GeV were simulated to determine the distribution of the number of ph.e. in the central ring, as well as the expected proton background rates. The solid line in Fig.5.9 shows this distribution for proton incident angles smaller than 5° . The level of the extragalactic LONS was included. The light associated with protons is comparable to the light

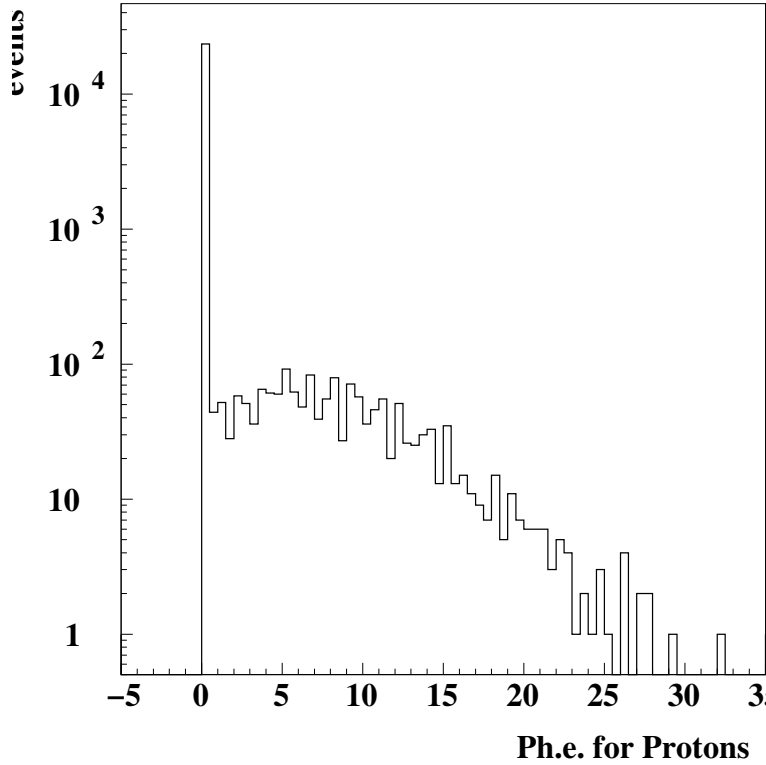


Figure 5.9: *Distribution of the number of ph.e. in the central ring due to proton showers (with a incident solid angle of less than 5°) plus diffuse night sky background. Note that the light produced by the protons is comparable to the light produced by low energy γ -ray showers.*

produced by γ -ray primaries at 5 GeV. However, we must also consider the influence of protons with energy from 2 TeV to 30 TeV. As a first approximation this effect can be estimated from rates obtained for the *normal* operation mode, but a proper simulation will be done in future work.

The rate of protons in the central ring was calculated including offset showers as discussed previously. The result is a rate of 30 Hz for 20 to 2000 GeV protons. Using the effective area given by de los Reyes et al., [32, de los Reyes, 2002] for the *normal* operation of MAGIC (see Fig. 2.8), the contribution of protons with energy between 2 and 30 TeV can be estimated: the average amount of light produced by a high energy proton shower is $R_{(2 \text{ to } 30 \text{ TeV})} = 1.16 \cdot R_{\leq 2 \text{ TeV}}$. The total proton rate will then be $\sim 65 \text{ Hz}$. An efficient rejection technique to reduce

the proton rate is discussed below.

- Cut in “size” parameter:

Size is defined as the sum of the number of ph.e in the shower image. We apply standard image cleaning procedure, i.e. we only select pixels that are 3σ above the noise level, and 2.5σ in a second level cleaning (tail cut and boundary cut). The total size is then:

$$S = \sum_i^N Q_i \quad (5.5)$$

where Q_i is the charge in the i^{th} PM and N the number of pixels surviving the trigger and cleaning conditions from the entire camera.

The quality factor (as defined in Chapter 2) associated to a cut in the size parameter can be optimized. Figure 5.11 plots the quality factor as a function of the size cut. Rejecting showers with size larger than $S=183$, we still keep 61% γ -ray showers and the proton rate will be 10% of 60 Hz which gives 6 Hz, which is acceptable given the high collection areas we have.

An approximation was made by simulating the normal 4NN trigger, from which the total shower sizes would be calculated as shown in Fig. 5.10 and Fig. 5.11.

- Cut in “distance” parameter:

Here we aim at using the rest of the camera as anti-coincidence. Proton initiated showers produce more extended images than γ -ray initiated ones. Therefore from those hadron showers which trigger the central ring, a large fraction will fulfill the *standard* 4NN condition in the outer ring of the camera whereas low energy γ -ray showers will be constrained to the inner

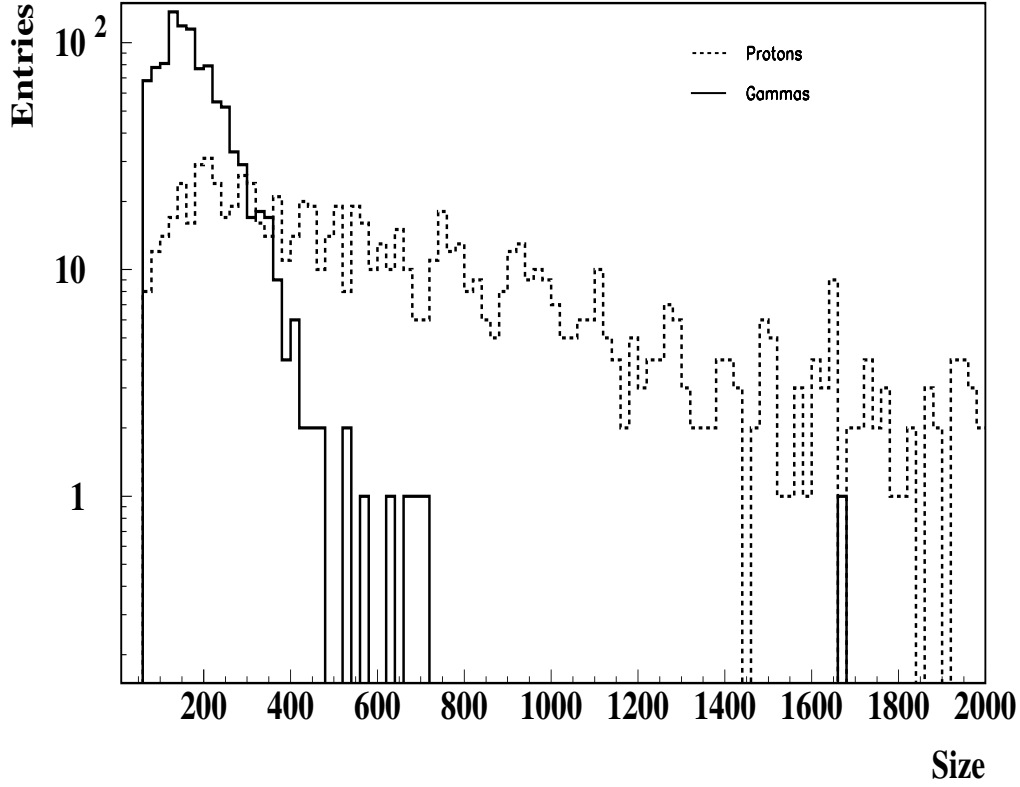


Figure 5.10: *Size distribution of γ -ray showers at low energy, from 5 to 10 GeV (solid line) and of proton showers from 30 GeV to 30 TeV (dashed line.)*

camera. An efficient proton rejection is expected using this pseudo-distance parameter.

A background rate $\leq 6 \text{ Hz}$ is expected to result from those off-showers which survive the size cut, but result in trigger at distances comparable to the radio of the camera. Such a cut will also reduce the γ -ray collection area by $\sim 40\%$, which implies a 40% reduction in the A_{eff} vs. ph.e.. curves in Fig.5.8.

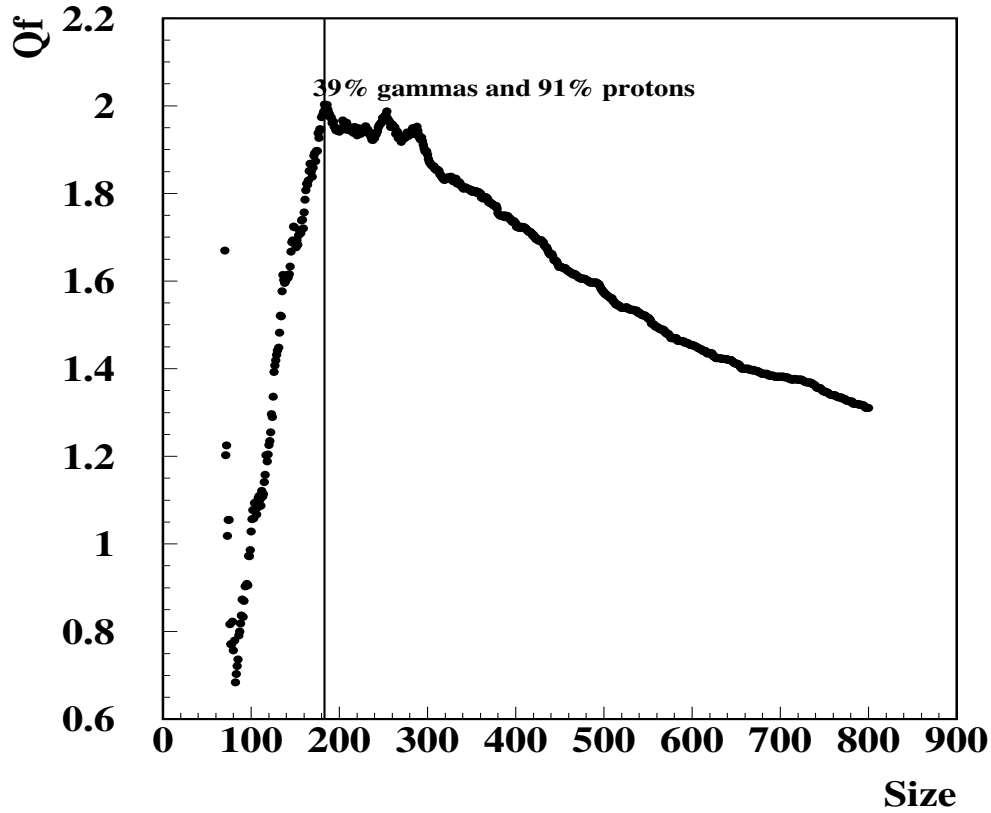


Figure 5.11: *Quality factor (Q_f) vs cut in Size (Size, in ph.e.)* The vertical line corresponds to a cut in size of 183 ph.e. that rejects 91% of the protons and 39% of the gammas.

5.3.2 Electron Background Rejection

ICATs have to deal as well with a background of electron initiated atmospheric showers. Showers initiated by electrons behave very much like γ -ray showers (see Section 1.1.2.) so electrons are practically indistinguishable from γ -rays. Like cosmic rays, primary electrons arrive at the Earth isotropically due to the deflections in the random component of the galactic magnetic field.

At arriving at our magnetosphere the Geomagnetic Field bends the electron/cosmic ray trajectories preventing low rigidity particles from reaching the atmosphere. The rigidity of a particle is defined as pc/Ze , where c is the speed

of the light, p is the momentum and Ze is the charge of the particle. Below a certain rigidity cutoff that depends on the telescope position at the Earth and the electron incident direction, the electron flux will be completely absorbed. The rigidity cutoff for electrons at La Palma has been calculated by Cortina & González [21, Cortina et al., 2001], and turns out to be ~ 10 GeV for the magnetic latitude of La Palma. That means that the electron background below 10 GeV can be neglected.

The expected rate for isotropic electrons above 10 GeV have also been calculated by Cortina & González for the MAGIC normal operation and the entire camera $R_{elec} = 3$ Hz. Taking into account that the area of the central ring is much smaller, together with the cut (discussed above) in size to reject showers with primary energy ≥ 10 GeV, a electron background rate $\ll 1$ Hz is expected.

Above 10 GeV, the rejection of showers with energy ≥ 10 GeV will be done with the size cut. Therefore, no electron background is expected.

5.4 Expected Rates

This technique is particularly useful when detecting transients such as neutron stars and AGNs. Some pulsars have a strong γ -ray component between 1 and 10 GeV, and above this range of energy a cutoff in the spectrum is expected [74, Nel & de Jager, 1995]. Another advantage is that an ON-OFF observation is not necessary since the OFF is taken when the variable source is not emitting, that means that the level of the LONS will be relatively constant during the observation time. Furthermore, if all 7 pixels from the central cluster consist of HPDs with optical monitors as discussed in Chapter 3, the LONS can be very accurately determined between ON and OFF observations of DC sources, in which case an accurate software padding will be possible [14, Cawley, 1993].

We know from the EGRET pulsar observations [95, Thompson et al., 1993] that γ -ray pulsar spectra can be fitted with a power law and an exponential

cutoff. The best fit for these parameters is represented in table 2.1, in Chapter 2.

The effective area is very dependent on the LONS level, as showed in fig. 5.8 and further studies must be done to determine the LONS level for each galactic pulsar to determine the detection sensitivity of each. The rate then is again calculated from the Eq. 2.20 in Chapter 2.

Therefore, for the 2 to 10 GeV energy range, and assuming a zenith angle of 0° (since our Monte Carlo simulation is only valid for vertical showers), we can estimate the γ -ray rates in the central ring, for a accidental LONS trigger rate of 0.1 Hz. For Crab, Geminga, the Crab nebula and a weak EGRET source the expected rate for detection by classical PMs are, using the scaling given in Fig.5.8 and the known pulsar fluxes above 1 GeV given in table 2.1:

LONS $ph \cdot m^{-2} s^{-1} sr^{-1}$	Crab pulsar Hz	Geminga pulsar Hz	Crab nebula Hz	Weak GeV source Hz
4.6×10^{12}	0.05	1.82	0.0011	0.005
1.7×10^{12}	0.20	6.40	0.044	0.25

Table 5.1: *Detection γ -ray rates for the Crab pulsar, Geminga pulsar, Crab nebula and a weak EGRET source assuming two different levels of LONS.*

Those rates will decrease in $\sim 40\%$ if a size cut is applied to reduce the cosmic ray background. But detailed simulations must be done once the camera program is modified to define a central ring trigger.

It is unlikely that Geminga will be detected using the *standard* operation mode (see Chapter 2) but it will show a high rate from 2 to 10 GeV if the LONS is at the level given by Mirzoyan et al. With a cutoff around 5 GeV and therefore maximum emission at this energy range, the Geminga pulsar should be considered as one of the best candidates to be observed by MAGIC with the central ring technique. As much as one photon per pulsar rotation is expected. The plerion nebula must also be taken into consideration. The Crab nebula has been already detected by EGRET and fitted by [29, de Jager, 2000] using the following spectrum:

$$\frac{dN}{dE} = kE^{-\Delta-c\text{Log}_{10}E}, \quad (5.6)$$

where the parameters are $k = 3.25 \times 10^{-11}$, $\Delta = 2.44$ and $c = 0.151$. The γ -ray rate expected at 2 to 10 GeV is also given in table 5.1.

To check in general the sensitivity of this technique, a weak EGRET source was selected from Lamb & Macomb, [58, Lamb & Macomb, 2000]. The γ -ray flux above 2 GeV (assuming a E^{-2} photon spectrum):

$$F_\gamma(\geq 2\text{GeV}) \approx 1 \times 10^{-4} m^{-2} s^{-1} \quad (5.7)$$

And the rates for two values of the LONS are given in table 5.1. Hence, if we assume that the size and distance cuts result in a background proton rate $Rate_{proton} \ll 60$ Hz and a γ -ray rate of $Rate_\gamma = F_\gamma A_\gamma$, the expression corresponding to a 5σ significance is:

$$\frac{Rate_\gamma T}{\sqrt{Rate_p T}} \geq 5 \implies Rate_\gamma \geq 5 \sqrt{\frac{Rate_p}{T}}, \quad (5.8)$$

a flux sensitivity of

$$F_\gamma(\geq 2\text{ GeV}) = \frac{0.08}{A_{eff}} \sqrt{\frac{Rate_p}{T}} \quad (5.9)$$

Therefore, depending on the LONS level:

$$F_\gamma(\geq 2\text{GeV}) = 2.5 \times 10^{-4} \left(\frac{Rate_p}{60\text{ Hz}}\right)^{\frac{1}{2}} \cdot T_{obs}^{-1/2} \quad (5.10)$$

$$F_\gamma(\geq 2\text{GeV}) = 1.2 \times 10^{-2} \left(\frac{Rate_p}{60\text{ Hz}}\right)^{\frac{1}{2}} \cdot T_{obs}^{-1/2} \quad (5.11)$$

where Eq. 5.10 corresponds to an average effective area of 2500 m^2 and Eq. 5.11

to an area of 50 m^2 . If a blue filter is applied, the sensitivity will be closer to Eq. 5.10.

Equating Eq. 5.7 with the expression for the effective flux above, (Eq. 5.9) it is possible to calculate the minimum observation time to detect a EGRET weak source using the MAGIC telescope. Thus, using an average effective collection area of 2500 m^2 , and a proton rate of $Rate_p = 60 \text{ Hz}$, the observation time required to detect an EGRET faint source would only be $T \ll 1 \text{ hour}$. This gives us an idea of the potential of this technique.

5.5 Geomagnetic effect in low energy showers

The Earth magnetic field plays a mayor role in the development of γ -ray showers at low energy, since the secondary electron/positron pair is separated by an angular spread $\Delta\theta \propto 1/E_e$. Thus the smaller the electron energy E_e , the larger the angular separation $\Delta\theta$ between the electron and positron and the larger the angular spread of the Cerenkov photons. The Lorentz force exerted in the charged secondary particles is proportional to the perpendicular component of the magnetic field, which has a magnitude of $B_\perp = 38 \text{ } \mu\text{T}$ at La Palma. The effect in the MAGIC imaging operation was investigated by de los Reyes et al. [33, de los Reyes, 2003]. We estimate its effect on the effective collection areas shown above, using the central ring technique. The perpendicular component of the magnetic field $B_\perp$ at La Palma for zenith angle less than 90° degrees is represented in figure 5.12. Zenith vs. azimuth angle is plotted and the different colors indicate the ratio of $B_\perp$ over its maximum value, with the red color the maximum effect, while the purple shows the minimum effect.

Three different zenith and azimuth angles have been selected to test the geomagnetic effect on 2 to 5 GeV showers. Their values and the corresponding effective collection areas are shown in table 5.2.

The reason for selecting zenith angle of 20° is because atmospheric absorption

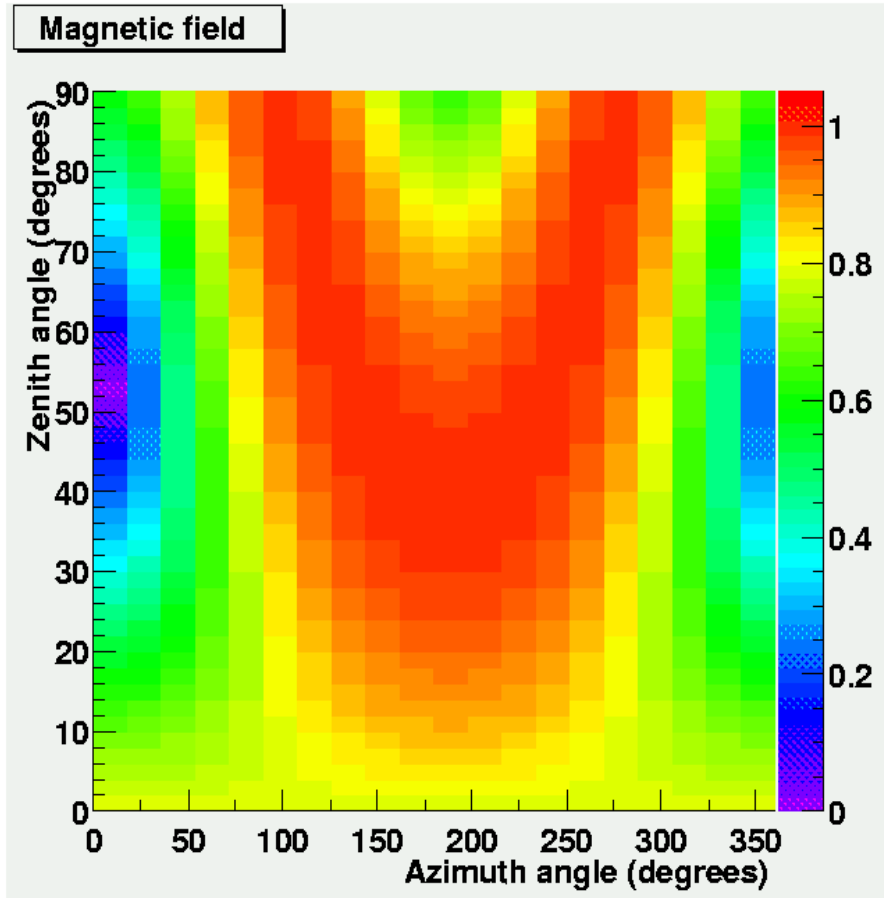


Figure 5.12: $B_{\perp}$ at La Palma latitude. The color contours give the ratio of $B_{\perp}$ over its maximum value. (Courtesy of R. de los Reyes)

will in any case reduce the collection area for 2 to 10 GeV showers if $z \gg 20^\circ$.

Azimuth angle	Zenith angle	Eff. Coll. Area(m^2)
0	20	2538
0	0	2097
180	20	1707

Table 5.2: *Effective Collection Areas for different zenith and azimuth angles.*

From Fig.5.2 it is clear that the geomagnetic effect ($B_{\perp}$) is larger at $z = 0^\circ$ (zenith), compared to $z = 20^\circ$ but Azimuth $Azim = 0$. However, for $z = 20^\circ$ and $Azim = 180^\circ$, we found that $B_{\perp}$ is again larger compared to the $z = 0^\circ$

case. Consequently the effective area is smaller where $B_{\perp}$ is at its maximum. On the contrary, when one goes to 180° azimuth and 20° zenith, the effect of the perpendicular magnetic field over the secondary particles is maximum so the effective area is reduced from 2538 to 1700 m^2 .

5.6 Conclusion

It is clear that we have exploited the large RMS in the photon density at low γ -ray energies to our advantage: whereas the RMS in the photoelectrons is comparable to the corresponding mean number of photoelectrons, the RMS for the LONS is only the square root of the mean. This feature allows us to capture the upward fluctuations of the photoelectrons arising from γ -rays, which are well above the tail of the LONS distribution. The fraction of such 5 GeV γ -rays that survive the selected trigger threshold is of the order of a percent, giving a collection area of 50 to 2500 m^2 (see Section 3) at 5 GeV, with the exact number depending on the LONS level.

In the case of the LONS associated with the Crab region, we find a collection area of 50 m^2 at 5 GeV (assuming that the central pixel containing the bright nebula was removed from the trigger. The expected γ -ray rates from the pulsar & nebula are then 5.1×10^{-2} and 0.1×10^{-2} Hz respectively, but can be increased to 0.20 and 0.04 respectively if the PMs are replaced by HPD (which have a larger QE), but covered by blue filters to reject LONS with $\lambda \geq 500$ nm.

With a background proton rate (ignoring information from the rest of the camera) from the central cluster of ~ 60 Hz, we expect a 3σ detection within ~ 60 hours for the Crab pulsar but the nebula will be very difficult to detect unless the LONS can be reduced with a filter as discussed above.

This new concept (even with limited spectral bandwidth) allows us to define similar 7-pixel clusters from adjacent pixels in the camera, with the normal direction to each cluster defining a new source direction, with the results of this

paper applying to each cluster. More simulations are required to establish the inter cluster spillover and hence final angular resolution.

This novel approach effectively defines a telescope with a total FOV of ~ 3.5 degrees and an angular resolution of ≥ 0.36 degrees. Apart from using the size cut to reject $\sim 90\%$ of the proton showers, which implies a maximum γ -ray energy of ~ 20 GeV, two or more coincident cluster triggers can also be rejected as proton background triggers, since preference for an upward fluctuation from a low energy γ -ray shower in a given cluster was already pre-selected, so that the probability for more than one cluster trigger by a γ -ray shower with energy below 10 GeV is relatively low. The camera outside the triggering cluster of 7 pixels therefore acts as an anticoincidence shield, with the fine pixelization allowing several possibilities to reject proton showers. This study with more details will be reported in a follow up paper.

Chapter 6

Discussion & Conclusion

In the work we have developed a good number of methods and ideas for pulsar studies with the MAGIC telescope, although most of the results can be applied to other γ -rays ground based telescopes. This thesis presents new principles towards the optimization of a 17 *m* reflector telescope at moderated altitude (2200 m) to detect pulsed γ -ray emission from neutron stars. Pulsars are believed to represent a large number of unidentified galactic sources in the EGRET catalogue, which covers up to 10 *GeV* energy but whose sensitivity and angular resolution above 1 GeV is not sufficient to do a positive identification of the γ -ray source on its error box.

It has been demonstrated in this dissertation that the MAGIC telescope, operative since October 2003, and with a large mirror reflector (17 *m* diameter) and technical improvements, will be able to detect pulsed emission from a few pulsars (canonical as well as millisecond pulsars) in a relative short time. The different theoretical models to explain pulsed γ -ray emission were discussed in Chapter 2. For millisecond pulsars, it was shown that the effect of the low magnetic field derived from observational parameters ($\sim 10^9$) shifts the typical cutoff in the energy spectrum to relative high energies. The condition for pair production on the perpendicular component of the electric field $E_\gamma B_\perp \geq 10^{18.6}$ must be fulfilled

to explain the turnover in the spectrum, but in the case of millisecond pulsar the magnetic field strength is too low to produce pair production. In absence or reduction of any pair production the curvature radiation would not be absorbed and the cutoff will be found at higher energies.

Applying these theoretical model conclusions based on general relativity, the spectra of a sample of millisecond pulsars have been calculated and it was found that the tail of the spectrum is extended sometimes up to TeV. Even when the γ -ray flux of millisecond pulsars is rather low, the large collection area ($\sim 10^5 m^2$) of the telescope at these energies ($\sim 30 GeV$ to a few TeVs) favors the detection. A total of 11 pulsars were selected as best candidates under the condition that, without any γ /hadron discrimination applied, are detectable in 48 hours. With the standard operation mode of the telescope, that requires 4NN trigger and 3.5 mV per pixel, from the millisecond pulsars detectable within 48 hours, the best candidates are PSR J1959+2048 and PSR 0030+0451. Not only for the short detection time required (15 and 12 h), but also for their relative latitude with respect to the MAGIC site, which avoid atmospheric extinction and hence the losing of sensitivity at large zenith angle. The model has been tested with Whipple observation of six millisecond pulsars and no contradiction was found with the upper flux delimited by their observations and the theoretical model.

The other large group of pulsars, the canonical pulsars, and particularly pulsars detected already by the EGRET satellite above 1 GeV, has been analysed to estimate the required observation time to achieve a detection with $\sim 3\sigma$, using the standard imaging technique (4NN trigger condition). For that purpose the effective collection area was calculated from Monte Carlo Simulation, either for showers initiated by γ -rays and protons. The Crab pulsar and PSR 1951+32 are the best candidates to be detected within 7.5 and 8.8 hours respectively. Their position relative to the MAGIC telescope site together with a high cutoff (30 and 40 GeV) makes the detection more feasible. The Geminga pulsar, the one with the lowest rate, due to the low cutoff, at around 5 GeV, will be very diffi-

cult to detect since its γ -ray flux drops superexponentially above this turnover. However, observation of γ -ray showers at very low energy will still be possible with the new technique proposed in Chapter 5. The imaging technique which has produced very successful results in the past, cannot be used anymore since at 2 to 10 GeV energy range, the pool of light becomes very faint and the LONS level in the whole camera makes these γ -ray showers indistinguishable. The technique we propose exploits the large reflector area of MAGIC to collect light and the fact that at 2 to 10 GeV, the fluctuation or RMS in the photon distribution are comparable to its average value. We restrict the FOV to 0.36° by using only the PMs in the inner part of the camera. The reason for this is because the amount of Cerenkov light inside this cluster of 7 pixels relative to the total light in the camera increase towards lower energies (see Fig. 5.4). The γ -ray detector now consists in the camera central pixel and the so-called central ring, a total of 7 pixels. Nevertheless, the results presented in this thesis do not include the high efficiency HPD in the central pixel. The remaining six PMs may be replaced by high efficiency HPD, in such a way that the sensitivity of the method will increase substantially.

A trigger threshold is selected over the sum of the currents coming from the 6 pixels that avoids most of the LONS accidental trigger. The fraction of such 5 GeV γ -rays surviving this selected trigger threshold is a small percentage, giving a collection area of 50 to 2500 m² at 5 GeV. However, this effective collection area is very dependent on the LONS level, further studies must be done to determine the real LONS around each pulsar or galactic source we plan to observe.

The LONS level in the central ring can be decreased reducing the angular resolution of the PMs, then the light from single bright stars can be eliminated by removing the brightest PM from the central ring trigger and only the contribution of the diffuse night sky background with fainter starlight remains. When removing one of the pixels the amount of Cerenkov light will decrease and so will the effective area. A optimization of the effective area versus LONS reduction

must be done before removing any PM. On the other hand, if a blue filter is applied, less than 50% of the Cerenkov light will be lost whereas the LONS level will be reduced in a factor 3 [71, Mineo et al., 2003]. The blue filter together with the high quantum efficiency HPDs will compensate the loss of light and effective areas of 2500 m^2 can be easily reached, for galactic sources with a high LONS level background, such as the Crab Pulsar. To determine the sensitivity of this technique, a EGRET weak source for energy $\geq 1 \text{ GeV}$, was selected, and it was found that, assuming an extragalactic LONS, $\ll 6 \text{ hours}$ is required to have a 5σ detection without any discrimination technique, while EGRET had typical observation times of 2 weeks. Geminga, which is the most unlikely detection with the normal operation mode, becomes a good candidate due to the high γ -ray emission at this energy range. Under the assumption that 7 HPDs are used in the central ring and a blue filter is applied, a rate of 6 Hz is predicted, which means 1 γ -ray per Geminga rotation.

Some proposals for γ /hadron discrimination were discussed as well.

- Using only the "size" parameter it is possible to reject 90% of proton showers and only 40% of γ showers. Given the high effective areas found, we can afford to loose this fraction of γ -rays without losing too much sensitivity since the expected rate below 10 GeV is still very high. This size cut can be applied as well to select low energy showers on the normal telescope operation mode (see Section 5.3.1).
- It is expected that proton showers that trigger in the central ring are the tail of the image of offset protons. Therefore we expect that those extended images will trigger in the outer part of the camera as well. Rejecting showers in this area, that is, a "pseudo distance" cut, we could reject a large fraction of protons. This technique can be applied comparing different clusters in the camera with the central cluster or the central cluster with cluster defined by the 4NN triggered pixels.

- The photon timing spread for low energy γ -rays is expected to be narrower than at higher energies since we only look at the first electrons/positrons. It is recommendable to apply faster FlashADCs (2 GHz) to optimize the 2 to 10 GeV γ -ray detection and reduce the LONS level. [63, Lorenz, 2003].

Further studies must be done to investigate the angular resolution of low energy showers. It is believed that when using the second ring around the central ring the LONS increases since the number of PMs increases and the efficiency of the technique decreases.

The effect of the magnetic field has been properly folded in, although not in too much detail. The magnetic field produces a deflection in the light due to its effect in electron/positron pairs in the shower. As expected, the effective area decreases as the perpendicular component of the magnetic field increases, and increases at certain azimuth and zenith angle when this effect is minimum. The detection power of this novel technique has been proved to be excellent although the energy resolution is rather poor.

Since Hillas parameters at low energy are not efficient anymore, it is possible to exploit periodicity statistical tests to detect transients such as pulsars. In Chapter 4 we discuss the timing analysis to fold into the γ -ray data, and a pulsar population study was investigated to determine the effect of timing noise and glitches in the extrapolation of radio ephemeris to the γ -ray observation data. A relation between the pulsar activity and its periodicity as well as its spin down power was found. As a conclusion it can be said that for most pulsars, extrapolation over 3 years does not suppose a problem for a few hours observation time. Nevertheless we must be careful when observing young pulsars with high spin down energy, such as the Crab pulsar. But for the Crab pulsar, beside the pulsar monthly ephemeris provided by the Jodrell Bank radiotelescope, the optical counterpart can be used to determine accurate parameters.

The Crab pulsar was detected in optical wavelengths using the 10 m^2 reflector

HEGRA-CT1 Cerenkov telescope, as a prototype to install a optical detector in the MAGIC central pixel. The Crab pulsed signal was detected within 200 sec. The sensitivity of the method using the increasing in the DC current was checked and the MAGIC sensitivity was also derived, using the same technique as in CT1 and also photon single counting. We concluded that MAGIC will be able to resolve single pulses if the proper technological improvements discussed in the text are introduced. The simultaneous optical/ γ observations have multiple applications. We specifically used the optical pulsed signal to determine the light of the night sky around the central pixel. A value of $I_{LONS} = 4.6 \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \cdot \text{s}^{-1} \cdot \text{sr}^{-1}$ was derived, which differs from the one calculated using a nearby star, ζ -Tauri, in 7%. This difference becomes smaller when considering the definition of main pulse and interpulse by Carramiñana et al. [13]. The total flux given by the sum of the pulse and interpulse does not cover the entire phase, so that there is a small contribution of the pulsed emission which was not considered. Therefore the value of the LONS calculated will decrease and the agreement with the value derived from ζ -Tauri measurements will improve. It is recommendable to measure the LONS for La Palma across the galactic plane for $\sim 0.4^\circ$ resolution ([71, Mineo et al., 2003]).

We showed that MAGIC will be able to do a lot in the γ -ray pulsar field. All this techniques will be tested very soon, and different theoretical models and physics incognitas will be solved.

Appendix A

List of Publications by E. de Oña Wilhelmi et al. relevant to this thesis.

- The Effect of Pulsar Timing Noise and Glitches on γ -ray Pulsar Observations. E. de Oña Wilhelmi et al. March 2002. African Skies/Cieux Africains 7.
- The Effect of Pulsar Timing Noise and Glitches on γ -ray Pulsar Observations II. E. de Oña Wilhelmi et al. Sept 2002 V Reunión Científica de la Sociedad Española de Astronomía. Space Science Library (ASSL Kluwer).
- The Effect of Pulsar Timing Noise on γ -ray Pulsar from Ground Based Gamma Telescopes Observations. E. de Oña Wilhelmi et al. August 2002, Radio Pulsar Conference in Crete. ASP
- The Effect of Pulsar Timing Noise and Glitches on Timing Analysis for Ground Based Telescope Observations. E. de Oña Wilhelmi et al. August 2003, the 28th International Cosmic Ray Conference.
- Determination of the Night Sky Background around the Crab Pulsar Using its Optical Pulsation. E. de Oña Wilhelmi et al. August 2003, the 28th International Cosmic Ray Conference.
- Detectability of γ -rays from Millisecond Pulsars with MAGIC. E. de Oña Wilhelmi et al. August 2003, the 28th International Cosmic Ray Conference.

- Determination of the Night Sky Background around the Crab pulsar using its optical pulsation. E. de Oña et al. Submitted to Astroparticle Physics, November 2003.

Appendix B

Due to a bug in the collaboration software we used, the results shown in Chapter 5 are not up to date. The results after correcting that bug are included in the article [79]. We include it in this appendix for completeness.

Ground-Based Pulsar Trigger Mechanisms For The 2-30 GeV Range

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Abstract. Large area ground-based γ -ray telescopes employ typically 3 to 4 nearest neighbor (NN) trigger mechanism, which makes them sensitive to γ -rays above 30 GeV, with the energy threshold depending on the area of the mirror and several other factors. In a novel new approach we investigate a complementary one, where a large central part of the imaging camera is placed in a fast summed trigger, while rejecting all NN events triggered by present systems. The result is a large collection area ($> 1000 \text{ m}^2$) in the 2-30 GeV range for a typical 17 meter class telescope, but somewhat larger energies for smaller sized telescopes. The resulting response function is ideal for new pulsar searches (timing analysis) of unidentified EGRET sources, with cutoff energies in the GeV range. New pulsars may be discovered with this technique, but at the expense of poorer angular and spectral resolution.

1. Introduction

It is well known that the Egret Instrument on the COMPTON Gamma-Ray Observatory had good background rejection capability as a result of its anti-coincidence shield, but it was rate limited as a result of its collection area of $\sim 0.1 \text{ m}^2$ up to about 30 GeV. The problem is that most GeV sources are weak, with threshold fluxes around a few times $10^{-4} \text{ m}^{-2} \text{ s}^{-1}$ above 1 GeV (Lamb & Macomb, 2000), so that typical space platforms, such as GLAST, with a collection area of $\sim 1 \text{ m}^2$, has to integrate for at least 10^4 sec to capture a few photons above a GeV from a region free of diffuse γ -ray emission.

This required integration time becomes at least 24 hours above 10 GeV. The latter constraint is however offset by the large sky coverage of GLAST



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(Dingus *et al.* , 1997). However, to see fast variability above GeV, much larger collection areas are required.

There is sufficient Physics motivation for observations above a few GeV: most canonical pulsars show spectral cutoffs in this range, whereas pulsar Electrodynamics predict hard spectra up to ~ 100 GeV from millisecond pulsars (Bulik *et al.* , 2000). Large redshift AGN observations are also not hampered by intergalactic absorption if photon energies down to a few GeV can be reached (Salamon & Stecker, 1998). Finally, the GeV range is also important for Particle Astrophysics (Flix *et al.* , 2003), (Blanch *et al.* , 2003), (López *et al.* , 2003).

To achieve the abovementioned goals with ground-based telescopes, an array of MAGIC type telescopes (17 m telescope)at an altitude of 5 km was suggested by Aharonian *et al.* (2001) with the purpose of achieving a large collection area at 5 GeV with good background rejection capability. Such an instrument would give excellent spectral response from a few GeV to multiple TeV energies, if the normal Imaging Atmospheric Cerenkov Technique (IACT)(Weekes *et al.* , 1989) is followed. Such a project however requires significant capital investment and infrastructure at high altitudes.

For the immediate future we offer a collection area of 800 to 2000 m² at 5 GeV, if the existing 17-meter class MAGIC Telescope (with its present $\sim 25 - 30\%$ quantum efficiency photon detector - (Cortina *et al.* , 2003)) is used.

The rationale, advantages and limitations are the following:

- (a) the angular resolution is limited to ≥ 0.50 degrees, with the latter radius defined by a cluster of 90 hexagonal pixels. This cluster defines a Cerenkov trigger (*pulsar trigger* or PT from now on), but consisting of those pixels without bright stellar content.
- (b) we show that the mean number of photoelectrons (ph.e) from a 5 GeV on-axis γ -ray shower is slightly larger than the mean number of the light of the night sky background (LONS) photoelectrons, if a 0.50 degree trigger radius at the center of the camera is selected.
- (c) for such a centralized cluster, we find that the rms of the number of photoelectrons induced by a 5 GeV γ -ray shower is comparable to the mean number of photoelectrons per shower, whereas the corresponding rms for the LONS is always equal to the square root of mean. Furthermore, the mean photoelectron level from such low energy showers is comparable to or slightly larger than the mean LONS level, which allows us to set a trigger threshold well above the mean LONS level, such that the accidental trigger rate on LONS events is negligible, whereas an appreciable trigger rate of 5 GeV showers can still be achieved.

- (d) the main background contribution arises from proton triggers inside the same PT area, which is about 260 Hz if any rejection technique is applied. We propose to reduce this rate to ~ 30 Hz if large events with high energies (>20 -30 GeV) are rejected, together with all coincident nearest neighbor (NN) triggers. In that way no events from any present telescope configuration survives, which includes events at higher energies and a high background level. This limits this special PT mode to EGRET energies, which is fine, since energies above 30 to 100 GeV is covered by the normal IACT, employed currently by the VERITAS, H.E.S.S., MAGIC & CANGAROO Collaborations, for which the normal 4 to 5 degree FOV pixelized imaging cameras were designed for image analysis as suggested originally by Weekes et al. (1989).
- (e) the collection area around 5 GeV is limited by the total LONS in the abovementioned cluster. Whereas a collection area of ~ 2000 m² may be reached at 5 GeV for the present MAGIC Telescope, a further factor 10 increase is achieved if the LONS is reduced by a factor of 3 on the average photo electron per 90 pixels. This demonstrates the necessity of removing at least the brightest pixels from the cluster in a dynamic feedback loop as discussed previously. From a Physics viewpoint, we expect extragalactic sources to yield the best sensitivity, given the lowest LONS level above the galactic plane.

The feasibility and sensibility of this technique and pulsed rates from known EGRET pulsar detection will be presented in the next sections.

2. Proposed Technique

For the PT technique, only the central part of the camera is used to detect very low energy γ -ray showers (< 10 GeV), exploiting the fact that low energy showers develop high in the atmosphere and therefore the pool of light falls mainly in the inner part of the camera.

The validity of this technique has been tried out with a telescope of the MAGIC telescope characteristics (Martinez *et al.*, 2003). We simulate the response PT of a MAGIC-like camera and reflector from 2 to 10 GeV showers and apply the technique. MAGIC is a 17 meter reflector telescope with a large field of view (FOV) pixelized camera ($\sim 3.7^\circ$). The camera detects Cerenkov light through 577 photomultipliers (PMs), each one with quantum efficiency of $QE \sim 25 - 30\%$, and mounted in a hexagonal frame.

The main problem to reduce the threshold of this type of ground based telescope is the high background (produced by the light of the night sky, LONS, and hadronic showers, mainly proton showers) which decreases the signal to noise ratio and makes difficult the identification of γ -ray showers

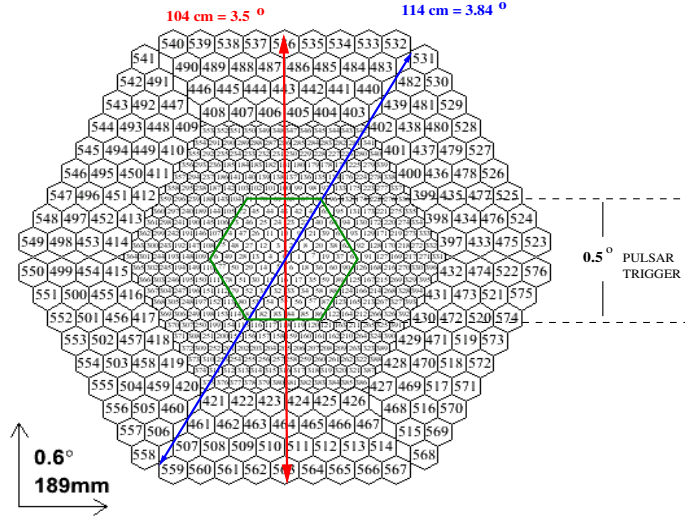


Figure 1. Schematic view of the MAGIC camera with 577 pixels. The inner camera is set up as 397 pixel of around 0.11° diameter while in the outer part the pixels size is $\sim 0.22^\circ$. The figure in the center shows the area in the camera which is for used in the PT mode.

versus hadron showers by *Imaging techniques*. Presently the MAGIC energy threshold is 60 GeV, being the foreseen threshold 30 GeV in phase I and ~ 10 GeV in phase II (with a camera of high quantum efficiency PMs).

Fig. 1 shows a schematic view of the dimension of the camera and the PT area. This PT area has been optimized to collect the maximum light from low energy showers and the best collection area above the LONS level. Therefore we restrict the light collectors to a 0.5° radius circle (PMs $\#i$, $i=1, \dots, 90$) in comparison with the 3.5° total camera FOV. This narrower angle of view allows us to reject part of the isotropic background. Moreover, showers at low energies hit the camera in this area. The summed signal from all pixels in the 0.5° radius will be used to detect the 2-30 GeV γ showers.

At hardware level, the analog signal detected for each PM must be sufficiently amplified to split it in two branches: one of them to be used for 2-30 GeV γ -ray detection (P_i in figure 2), and the other for the *standard* MAGIC operation mode (N_i). A proper impedance matching must be implemented to avoid pulse reflections in the camera between the NN mode and PT. Outputs given by N_1 to N_{90} for the PT area, combined with the pixel information from N_{91} to N_{577} are used to estimate usual Hillas parameters (Weekes *et al.*, 1989). This information can be used as anti-coincidence technique to reject both high energy γ -ray events and the high hadronic background associated with such type of triggers.

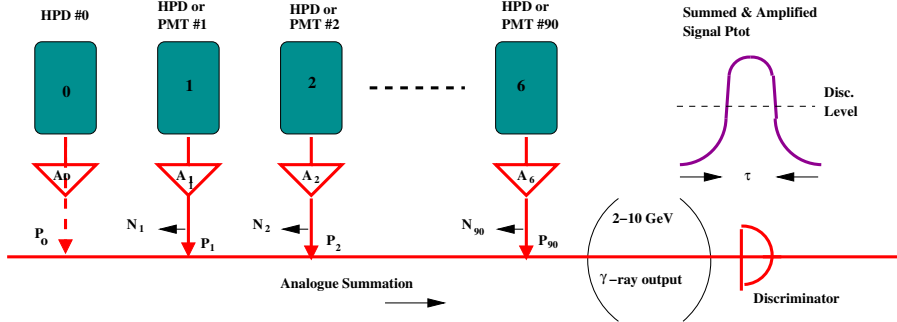


Figure 2. Pixels (PMs or higher efficiency hybrid photo detectors, HPDs) within the PT area P_i $i=1, \dots, 90$ are drawn with a schematic view of the electronic configuration. The central pixel, empty at the moment but to be used in optical studies, is expected to be included in future. The signals from the PMs will be summed P_{tot} and a discriminator sets the trigger level to allow an acceptable LONS accidental trigger rate.

The threshold imposed by the discriminator will depend on the level of the LONS. Depending on the source position on the sky, in a more or less noise background, we will allow different LONS accidental trigger rates. The trigger gate time should be shorter than in the NN operation mode $\tau=3$ ns (or less), since the characteristic time spread of low energy γ -ray showers is expected to be narrower due to their shorter path in the atmosphere.

3. Simulation Studies

A set of 2×10^5 γ -ray showers between 2 and 10 GeV and 10^5 protons in a 1 GeV to 30 TeV energy range, with the canonical cosmic rays spectrum (slope=-2.7) was simulated. The air showers in the atmosphere are calculated with the MMCS (MAGIC Monte Carlo Simulation) based in the CORSIKA program (Knapp & Heck, 1998). It consists of a detailed Monte Carlo program to study the evolution of EAS in the atmosphere initiated by photons, protons, nuclei, or any other particle. All particles are tracked along their trajectories and the output of the program gives type, energy, location, direction and arrival times of all secondary particles that are created in an air shower and pass a selected observation level. The simulation of the response of the MAGIC telescope is done with two programs developed by the MAGIC collaboration. The reflector program (Bigongiari *et al.*, 2003) is using the output files with the Cerenkov particles of MMCS as its input together with the geometrical information about the mirrors. The camera program is at the end of the chain (Blanch & Moralejo, 2004). This program is simulating the behavior of the PMs and of the electronics of the MAGIC camera. The camera program was modified to introduce the PT condition explained above.

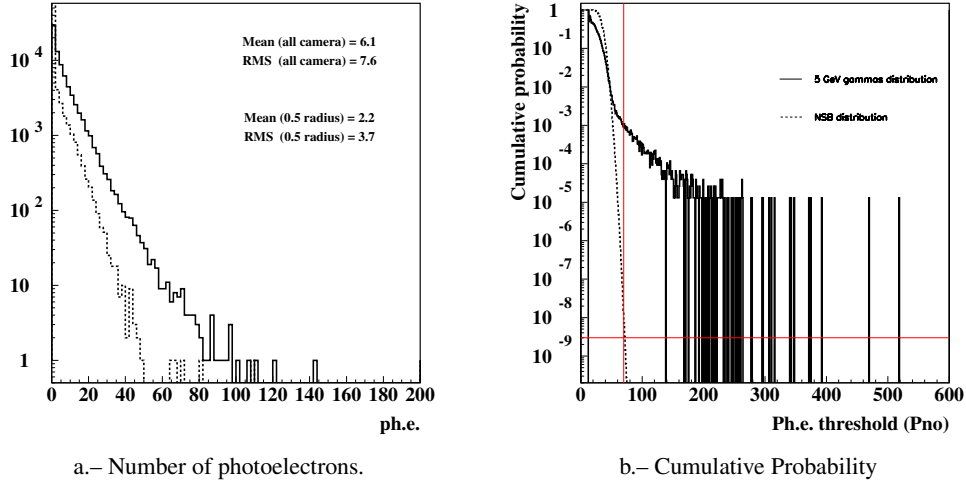


Figure 3. On the left (a) Distribution of the number of ph.e. generated in the camera from γ -ray showers with 5 GeV initial energy. The number of ph.e. from the PT area is represented with a dotted line. A mean value of ~ 6 ph.e. means a Čerenkov photon density from the shower of $0.2 \text{ ph}/\text{m}^2$. A 30% of these photons will be collected in a 0.5° radius around the center of the camera (15% of the total camera area). On the right (b) Distribution of number of ph.e. plus diffuse night sky background for 5 GeV showers. The rms of the signal is very similar to the mean value, therefore, the number of ph.e. that survives after threshold is still quite high. The poissonian distribution (dotted line) for a LONS intensity corresponding to an extragalactic source is also included in the figure. The horizontal line marks a LONS accidental trigger rate of 1 Hz, corresponding to a probability of 10^{-9} . It defines a trigger threshold of 70 ph.e., marked with a vertical dotted line.

Fig. 3(a) shows the number of photoelectrons for γ -ray showers at 5 GeV initial energy in the total camera. The distribution of the number of ph.e. collected in the PT area is drawn with dotted line the same figure. The mean value of the number of ph.e. in the total camera is ~ 6.1 . This value corresponds, scaling with the mirror reflectivity (Ref=85%), PMs quantum efficiencies and telescope area ($A_{MAGIC} = 234 \text{ m}^2$), to a density of photons of $\sim 0.2 \text{ ph} \cdot \text{m}^{-2}$. The mean value of the photon density agrees with the value already estimated by other authors (Konopelko, 1997). In both distributions (ph.e. in the total camera and in the 0.5° radius area), the rms is comparable to the mean value. This feature of low energy showers will lead us to relatively large trigger efficiency and therefore large collection area.

The total summed signal in the PT area P_{tot} , including the value of a extragalactic LONS (Mirzoyan & Lorenz, 1994), is recorded for each shower. Fig. 3(b) shows the cumulative probability of the number of ph.e. plus LONS value versus a certain threshold over the number of ph.e., P_{no} . The poissonian

probability for the extragalactic LONS background which was added to the signal is also represented with dotted line. Using the intensity given by (Mirzoyan & Lorenz, 1994): $I_{LONS}=1.7 \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \text{s}^{-1} \text{sr}^{-1}$, the LONS rate λ_{LONS} in $\tau = 3 \text{ ns}$ per pixel, which subtends a solid angel Ω , is:

$$\lambda_{LONS} = I_{LONS} \cdot A_{MAGIC} \cdot Ref \cdot QE \cdot \Omega \cdot \tau = 0.36 \text{ ph.e} \quad (1)$$

Therefore the total contribution in a 0.5° radius circle, which includes 90 photo detectors, follows a poissonian distribution centered at $\sim 33.5 \text{ ph.e}$.

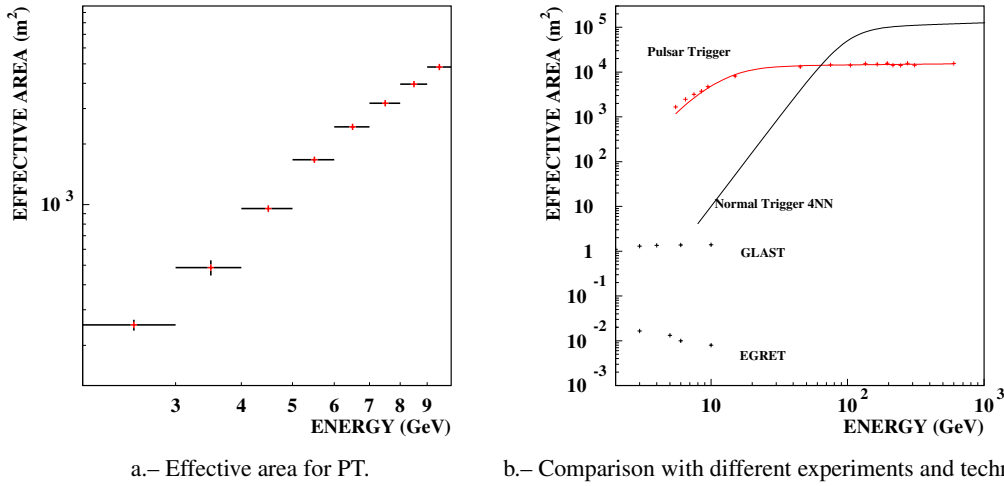


Figure 4. On the left (a) Effective area between 2 and 10 GeV using the PT technique. A mean value of 2000 m^2 is reached allowing a LONS accidental trigger of 1 Hz. On the right (b) Comparison of the effective area for two different methods for the MAGIC telescope, using the PT mode, and NN imaging technique. The effective areas of the two main satellites in the same range of energies, EGRET and GLAST, are also shown.

The selection of a trigger threshold, which avoids most of the accidental triggers is required. If T_n is the number of events which induces more than T_{no} ph.e. in the PT area, the accidental trigger rate will be given by rate B_{LONS} :

$$B_{LONS} = \frac{\tau}{\text{Prob}(T_n \geq T_{no})} \quad , \quad \text{Prob}(T_n \geq T_{no}) = \sum_{i=T_{no}}^{\infty} \frac{e^{-\lambda_{LONS}} \lambda_{LONS}^i}{i!} \quad (2)$$

as a function of the well known poissonian probability.

For different LONS accidental rates (1 Hz, 0.1 Hz), the arrival times and pulse height P_{tot} will be recordered if the trigger condition is fulfilled. Once the trigger condition is determined, the effective collection area (Fig. 4(a)) in the 2 to 10 GeV energy range can be derived.

A mean value of 2000 m^2 for the effective area is reached with this new technique for extragalactic γ -ray sources. This effective area can be compared with the effective area using the standard trigger condition NN (4 ph.e. and 4 NN pixels). Fig. 4(b) shows the effective area of the MAGIC telescope versus energy, using the PT mode from 2 to 3000 GeV, and the NN operation mode (de los Reyes *et al.*, 2002). The effective collection area of other γ -ray experiments are marked in the figure as well: EGRET, with a mean effective area in the 1 to 10 GeV range of $\sim 10^{-2} \text{ m}^2$ (Thompson *et al.*, 1993) and GLAST, which is expected to reach $\sim 1 \text{ m}^2$ above 1 GeV. An analytical expression of the effective area as a function of energy, for energies less than 20 GeV, can be calculated for the 2 levels of LONS considered previously:

$$A(E, I_{\text{extragalactic LONS}}) = 69 \cdot \left(\frac{E}{1 \text{ GeV}}\right)^{1.9} \text{ m}^2 \quad (3a)$$

$$A(E, I_{\text{galactic LONS}}) = 14 \cdot \left(\frac{E}{1 \text{ GeV}}\right)^{2.4} \text{ m}^2 \quad (3b)$$

4. Background Rejection

The main contribution for the background when detecting γ -ray showers with energies below 10 GeV with a ground based telescope, comes from the LONS and from hadronic showers. In the NN operational mode the imaging technique has been successfully proven, but for the PT technique a new approach must be considered.

The PT technique is highly sensitive to the LONS level, that is, to the source position. Fig. 5(a) shows this dependency for 5 GeV γ -ray showers, with a LONS accidental trigger of 0.1 Hz. The effective area becomes optimal for extragalactic sources such as 3c279 (see next section) while for galactic pulsars such as Crab it decreases below 1000 m^2 . On the other hand, canonical pulsars show maximum emission at a few GeV so this fact offset the smaller effective area.

Regarding the cosmic background, the main contribution comes from proton showers (96% of cosmic rays). The background rate has been calculated for a set of proton initiated showers with energies from 1 GeV to 30 TeV (although below 30 GeV the light produced is negligible), to be 260 Hz. In a first approximation, no imaging technique can be applied (although some patterns with less restricted conditions than NN are being investigated), and the protons rejection becomes more complicated. Nevertheless, the information given by the standard operation mode will allow us to reduce the proton background to acceptable levels. All events that fulfill 4NN conditions can be rejected. Those events also trigger in the less restricted PT condition, in such

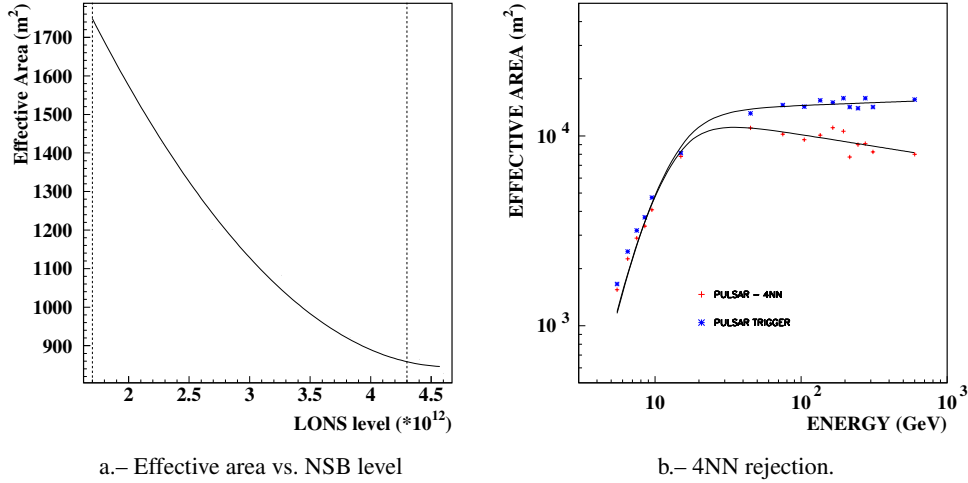


Figure 5. On the left (a) Effective area for the MAGIC telescope at 5 GeV using the PT technique vs. the LONS background level, allowing 0.1 Hz accidental trigger. The two vertical lines mark the 2 measured values, for a extragalactic source and for the Crab region in the galactic plane (Oña-Wilhelmi *et al.*, 2003), $I_{LONS} = 4.3 \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \text{s}^{-1} \text{sr}^{-1}$. On the right (b) Effective area for PT vs. energy. The effective area using only the PT condition is marked with stars, while the crosses mark the results if 4NN events are rejected.

a way that, if the rate of protons resulting from the 4NN trigger condition is 200 Hz (de los Reyes *et al.*, 2002), rejecting those events the total rate will be $\sim 60 \text{ Hz}$. On the other hand, low energy γ -ray events do not survive the 4NN condition, as discussed above. Therefore, by rejecting all 4NN trigger events, the PT efficiency above 20 GeV decreases while it holds up for lower energies. Fig. 5(b) shows this effect of combining the 2 triggers together. In the next section we will discuss the expected rate with and without reducing the background rate.

5. Expected Rates

This technique is particularly useful when detecting transients such as neutron stars and extragalactic sources. The timing analysis can overcome the difficulties due to a high LONS or proton background. Moreover, most of canonical pulsars have a strong γ -ray emission between 1 and 10 GeV, while, above this range of energy, a cutoff in the spectrum is expected (Nel & de Jager, 1995). Another advantage is the stability of the LONS level since OFF observation may be taken during the off-pulsation. The pulsed emission detected by EGRET (Thompson *et al.*, 1993) from several neutron stars can be

fitted with a power law and an exponential cutoff:

$$\frac{dN_\gamma}{dE} = K E^{-\Gamma} e^{-(E/E_o)^\alpha} \quad (4)$$

where Γ is the photon spectral index in the highest energy range detected by CGRO. The cutoff energy is given by E_o , whereas α determines the strength of the cutoff.

Table I shows the best fit for these parameters.

Table I. *Table of parameter fits for eq 4*

Pulsar	$k (\times 10^{-8})$ ($cm^{-2}s^{-1}GeV^{-1}$)	Γ	E_o (GeV)	α
Crab	24.0	2.08	30	2
Vela	138	1.62	8.0	1.7
Geminga	73.0	1.42	5.0	2.2
PSR B1951+32	3.80	1.74	40	2
PSR B1706-44	20.5	2.10	40	2

To calculate the expected rate for these pulsars it is necessary to use a lower effective area for 4.6×10^{12} LONS level. This γ -ray rate was calculated assuming the sources at 0° zenith angle.

Finally to show the potential of this technique with extragalactic sources, the blazar 3c279 detected above 1 GeV with EGRET (Hartman *et al.*, 1992) has also been considered, with a differential flux:

$$\frac{dN}{dE} = (2.7 \pm 0.3) \times 10^{-7} E^{2.02 \pm 0.07} cm^{-2}s^{-1}GeV^{-1} \quad (5)$$

The expected rates of the sources discussed above are presented in table II, using a background rate of 60 Hz.

Most of the pulsars can be detected in less than 30 min, if 4NN events are vetoed for our analysis. In the case of 3c279, with a redshift of 0.55 we can use a less restrictive condition for the trigger level due to the lower LONS level. The technique then is optimal and the high rate above 1 GeV allows us to detect it within less than 20 min. This source is one of the best candidates.

Selecting a weak EGRET source (Lamb & Macomb, 2000), it is possible to check the sensitivity of the PT technique. The γ -ray flux of a typically weak source above 1 GeV is

$$F_\gamma(\geq 2GeV) = 2.5 \times 10^{-4} m^{-2}s^{-1} \quad (6)$$

Table II. *Expected rates for different sources detected by EGRET below 10 GeV, using the PT technique.*

Source	Rate (Hz)	T_{obs} (min)
Crab	2.2	5
Vela	5.5	< 1
Geminga	1.7	9
PSR B1951+32	1.1	20
PSR B1706-44	2.3	5
3c279	290	< 0.1

if 5σ is required for a positive detection:

$$\frac{Rate_{\gamma}T}{\sqrt{Rate_p T}} \geq 5 \implies Rate_{\gamma} \geq 5\sqrt{\frac{Rate_p}{T}} \quad (7)$$

Hence, with T, observation time in hours, the effective flux (with $Rate_{\gamma}=F_{\gamma}A_{eff}$):

$$F_{\gamma}(\geq 2 \text{ GeV}) = 2.5 \times 10^{-4} \left(\frac{2500 \text{ m}^2}{A_{eff}} \right) \left(\frac{Rate_p}{60 \text{ Hz}} \right)^{1/2} \left(\frac{1 \text{ hr}}{T} \right)^{1/2} \text{ m}^{-2} \text{ s}^{-1} \quad (8)$$

which is comparable to the typical fluxes of weak GeV EGRET sources (for T=1 hr), provided that the spectrum extends up to at least 10 GeV. Equation 8 was normalised to an average effective area of 2500 m^2 for the 2 to 10 GeV range, assuming a LONS corresponding to the galactic plane anticenter (near the Crab pulsar).

6. Discussion

This new technique is based in the exploitation of the large rms in the photon density at low γ -ray energies to our advantage: whereas the rms in the photoelectrons is comparable to the corresponding mean number of photoelectrons, the rms for the LONS is only the square root of the mean. This feature allows us to capture the upward fluctuations of the photoelectrons arising from γ -rays, which are well above the tail of the LONS distribution. The fraction of such 5 GeV γ -rays surviving the selected trigger threshold is ~ 0.2 to 2%, giving a collection area of 1000 to 4000 m^2 (see Section 3) at 5 GeV.

The results presented in this paper have been done assuming an ideal detector. In a realistic implementation, detector effects will introduce unwanted noises and phenomena such as after pulsing, which may require modifications in the proposed setups.

The sensitivity of the method to the LONS level makes it easier to detect extragalactic sources, such as the quasar 3c279, which may be detectable in less than 10 sec. Thus, despite the higher LONS level expected for the galactic place (compared to high galactic latitudes), we should be able to detect existing EGRET pulsars, as well as new ones by exploiting the timing signatures of young and middle aged pulsars. This technique should be stable, since the pulsar periods are much shorter than the timescales for gradients in the LONS to reflect as changes in the instrument sensitivity and accidental trigger rate. The fact that we should be able to make a detection within a few hours, will naturally result in significantly reduced statistical trials when searching for new pulsars.

If we manage to reduce the LONS level better results are expected. The detection power of this novel technique has been proven to be excellent although the energy and angular resolution is rather poor. Studies about the effect of the magnetic field will be reported in a follow up paper.

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Appendix C

Determination of the Night Sky Background around the Crab pulsar using its optical pulsation

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Abstract

The poor angular resolution of imaging γ -ray telescopes is offset by the large reflector areas of next generation telescopes such as MAGIC (17 m diameter), which makes the study of optical emission associated with some γ -ray sources feasible. Furthermore, the extremely fast time response of photomultipliers (PMs) makes them ideal detectors for fast (subsecond) optical transients and periodic sources like pulsars. The optical pulse of the Crab pulsar was detected with the HEGRA CT1 central pixel using a modified PM, similar to the future MAGIC camera PMs.

The HEGRA CT1 telescope is a 2.1 m radius Čerenkov telescope, which has seen the Crab optical pulsations. The corresponding observation time required for a detection is 200 seconds, but MAGIC will only require ~ 30 sec if the same strategy with standard PMs is used.

The constancy of the pulsed fraction (with a relatively small error) during the 2 hours CT1 measurements shows that the pointing did not change during the

observations. The purpose of these periodic observations was to determine the light of the night sky (LONS) for the galactic anticenter Crab region, and to compare it with the LONS for a nearby bright star (ζ Tauri). These obtained LONS values vary between 4.6 and $4.3 \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \cdot \text{s}^{-1} \cdot \text{sr}^{-1}$ (with a coarse estimated error of 10%), being the first value from the timing signature of the Crab pulsar, and the second one was derived from the technique using the ζ Tauri standard star. Our results are between 2.5 and 3 times larger than the previously measured LONS (outside the galactic plane) of $(1.7 \pm 0.4) \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \cdot \text{s}^{-1} \cdot \text{sr}^{-1}$, as expected since the Crab pulsar is in the galactic plane, which implies a slightly higher energy threshold for Crab observations, if the higher value of CT1 measured LONS rate for galactic sources is used.

Key words:

Crab Pulsar, Night Sky Background, γ -ray telescope

PACS: 95.55.Ev, 95.55.Ka, 97.60.Jd

1 Introduction

Imaging Atmospheric Čerenkov Telescopes (IACTs) can be used to detect the optical emission of an astronomical object through the increased DC current of the camera pixels. Each of these pixels has a field-of-view (FOV) in the range of 0.1-0.3°. Hence, when an IACT follows a star, its optical emission is fully contained in the telescope central pixel.

The Crab Nebula is known to be the most important calibration source for

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high energy astrophysics, and especially for the ground-based atmospheric IACTs. The night sky background (LONS) influences the trigger threshold, the energy and the spectral determination of γ -rays, therefore the LONS around the Crab Nebula must be accurately known in the wavelength interval of the photomultipliers (PMs) used for the IACT cameras. Mirzoyan and Lorenz (7) measured the LONS at the site of the HEGRA experiment for a narrow angle, using a detector attached to the telescope CT2, which integrates only the starlight from the background of faint stars, obtaining a value of $I_{LONS} = (1.7 \pm 0.4) \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \cdot \text{s}^{-1} \cdot \text{sr}^{-1}$ (outside the galactic plane). In this paper we measure this I_{LONS} for the Crab region and we compare it with the value of (7), using a modified PM in the central pixel of the CT1 camera.

The optical pulsed emission of the Crab pulsar has already been detected using IACTs (9) and other Čerenkov detectors (4). The main purpose of such studies are to determine the timing parameters of the pulsar, with which the γ -ray arrival time can be folded, thus reducing the search for pulsed emission to a single statistical trial. It also allows us to test the accuracy of the used timing system and the solar system barycentric correction software, when comparing the timing parameters against contemporary radio data. This paper reports the optical timing signature of the Crab pulsar obtained, using the central pixel of CT1 (8), and the DC observations of the bright nearby star ζ Tauri, used to derive the LONS around the Crab Nebula. These two values cross-check the LONS measurements obtained using two different techniques: temporal (Crab pulsation) versus spatial (nearby star).

Our results with CT1 can be extended to a much larger telescope such as MAGIC (1) (17 meter diameter). IACTs are calibrated using as reference the Crab signal, and we can therefore use this regular Crab monitoring, not only

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for Čerenkov measurements, but also to measure the optical Crab pulsation with a modified PM in the telescope camera. Furthermore, since the astronomical community does not monitor the Crab pulsar continuously (on a 24 hour timescale) in radio or optical, observations with Čerenkov telescopes will also help to serendipitously discover pulsar glitches or period irregularities.

2 The Crab Optical Spectrum and Expected Single Photoelectron Response.

The optical spectra of the Crab pulsar main pulse and interpulse, and the spectrum of the underlying nebula (2), follow the power laws given in table 1.

The nebula with a size of $4' \times 6'$ fits well into a single pixel.

Flux	K	α
Main pulse	5.9×10^{-15}	$+0.2 \pm 0.1$
Interpulse	1.9×10^{-15}	$+0.2 \pm 0.1$
Underlying nebula	8×10^{-12}	-0.4 ± 0.1

Table 1

Flux in the 0.1 phase window centered on the main pulse and interpulse, where, $F_\lambda = K \times (\lambda/\lambda_o)^{-\alpha-2}$, $\lambda_o = 6000 \text{ \AA}$, $5000 \text{ \AA} < \lambda < 7500 \text{ \AA}$. The useful wavelength range is 3000-8000 $\text{\AA}$. Fluxes are in $\text{erg} \cdot \text{s}^{-1} \cdot \text{\AA}^{-1} \cdot \text{cm}^{-2}$.

We can calculate the single photoelectron rate for CT1 and MAGIC. Note that the poor angular resolution (a single pixel with FOV of 0.25° for CT1 and 0.1° for MAGIC), is compensated for by the large mirror area A (10 m^2 for CT1, 250 m^2 for MAGIC).

The expected photoelectron rates are calculated with the following expression:

$$R = \int F_{\lambda} \cdot \eta \cdot Q_{\nu} \cdot A \cdot \beta \, d\nu, \quad (1)$$

where η is the mirror reflectivity and Q_{ν} the frequency dependent quantum efficiency of the PMTs. The parameter β takes into account the photoelectron detection on the first dynode (0.9) and the light guide efficiency (90%).

For the Crab pulsar, the theoretically expected rates (Eq. 1) are given in Table 2, for CT1 and MAGIC telescopes.

Telescope	Total Pulsar Rate (phe $\cdot s^{-1}$)	Nebula Rate (phe $\cdot s^{-1}$)	LONS Rate (phe $\cdot s^{-1}$)	Background Rate (phe $\cdot s^{-1}$)
CT1	1.5×10^5	1.4×10^8	3.3×10^7	1.7×10^8
MAGIC	3.7×10^6	3.5×10^9	1.3×10^8	3.6×10^9

Table 2

Expected currents (in photoelectrons/s) of the optical emission of the Crab pulsar in the central pixel of the CT1 and MAGIC telescopes. Background rate means the total background: LONS + Nebula

If we assume that the PM has no inherent fluctuations (for example the fluctuations of the electron multiplication process between dynodes), we may derive the basic scaling parameter $X = p\sqrt{(N)}$ based on photon (Poissonian) statistics, for all test statistics, which measures the integrated difference between the measured pulse profile and a uniform distribution (no pulsed signal) (3). The parameter p is the number of pulsed events, divided by the total number of events N (signal plus noise). This X also corresponds approximately to the expected DC excess (measured in Gaussian standard deviations). Assuming that the ephemeris for folding the arrival times of optical photons from the Crab pulsar to phases is correct, the expected scaling parameters for CT1 and

MAGIC (taking I_{LONS} from (7), although we expect it to be higher), based on photon statistics alone, are:

$$X_{ct1} = 11\sqrt{\frac{T}{1 \text{ sec}}} \quad (2)$$

$$X_{MAGIC} = 62\sqrt{\frac{T}{1 \text{ sec}}} \quad (3)$$

From the basic scaling parameter we infer the expected time to detect the optical pulsation for a given significance. The theoretical rate from the Crab pulsar shows that it would be possible to detect the Crab optical pulse with the CT1 telescope in $T=75$ ms, with a significance of 3σ . For MAGIC, the situation is even better because the main pulse could be detected within 5 ms, which is the approximate duration of this pulse. This implies that MAGIC has the capability to monitor single pulses from Crab, based on theoretical arguments, and using the LONS intensity corresponding to an extragalactic source.

This result is only valid if the photodetector itself has no noise, and the only source of fluctuations is Poissonian in nature, due to the signal. The expected detection times increase dramatically when we take fluctuations into account, as shown by the measurements presented below.

3 Measurements

Measurements were done with the HEGRA CT1 telescope, during February 11-16th 2002 and during November 1-8th 2002. The standalone HEGRA CT1 telescope has a reflector area of 10 m^2 and a camera consisting of 127 PMs, each with 0.25° FOV.

Datasets were obtained with the Crab in the central pixel, and OFF source runs were also taken to check for any systematics. For optical observations the central pixel PM was modified. The DC branch, designed to monitor the DC current of the pixel, was adjusted to detect pulses of ~ 3 ms (timescale of the pulse width), and the AC branch, which is designed to transmit the ns fast signals generated by the Čerenkov showers was removed (5).

The DC current was measured using a National Instruments PCI NI 6043E digitising card. The sampling rate was set to 2 kHz and the pixel line was AC-coupled to reject low frequencies, and increase the card's effective dynamic range.

4 Analysis and Results of the Pulsed Optical Signal Measurements

The arrival times t_i were folded to obtain the phases Φ_i .

$$\Phi_i = \Phi_0 + \nu(t_i - t_0) + \frac{1}{2}\dot{\nu}(t_i - t_0)^2, \quad (4)$$

where $\nu, \dot{\nu}$ are the frequency and frequency derivative, and t_0 is the reference epoch.

To transform these arrival times to the solar barycentre system, we used the TEMPO software (6). A correction due to a small average drift in the digitisation card clock, as measured with respect to an atomic reference clock, was also introduced. The Jodrell Bank ephemeris (frequency, first frequency derivative and reference time) applicable to these observations are given in Table 3 (Jodrell Bank Crab pulsar monthly ephemeris).

Date	t_o (MJD)	ν (Hz)	$\dot{\nu}$ ($10^{-15} \text{ sec}^{-2}$)
15th October 2002	52562	29.8132421622	-373777.06
15th November 2002	52593	29.8122410972	-373758.90

Table 3

Radio Timing Parameters for the Crab pulsar for the indicated epochs.

A phaseogram was produced for each independent frequency confined to the wide frequency range between 26.0 and 32.0 Hz. The independent Fourier spacing (IFS) between this range of independent frequencies is known to be $1/T$ (periodogram bin size), where T hours is the observation time (2 hours).

The folded intensities for each test frequency were tested against a uniform distribution by performing a χ^2 fit to a constant intensity. The reduced χ^2 was calculated (Figure 1). A maximum value of the reduced χ^2 of ~ 154 was found and the periodogram signal is contained within one IFS of the expected Jodrell Bank ephemeris, consistent with the uncertainty in the drift expected from the clock on the digitisation card. Absolute phase matching with the expected radio/optical/X-ray/ γ -ray main pulse position was not possible, since an absolute UTC match was not done, and the drift was not accurately measured during observations.

Furthermore, the frequency resolution for the 2 hour observation was too coarse to resolve the differences between the ephemerides corresponding to 15 October and 15 November. The predicted frequency shown in Figure 1 is therefore consistent with both ephemerides.

Figure 1 shows our measured Crab pulse profile, with two sharp peaks, with a ~ 0.42 phase difference. The corresponding pulsed fraction can now be used

to isolate the LONS and cross-check the previous measurement (7). Using our results we can infer the MAGIC response for the Crab pulsar.

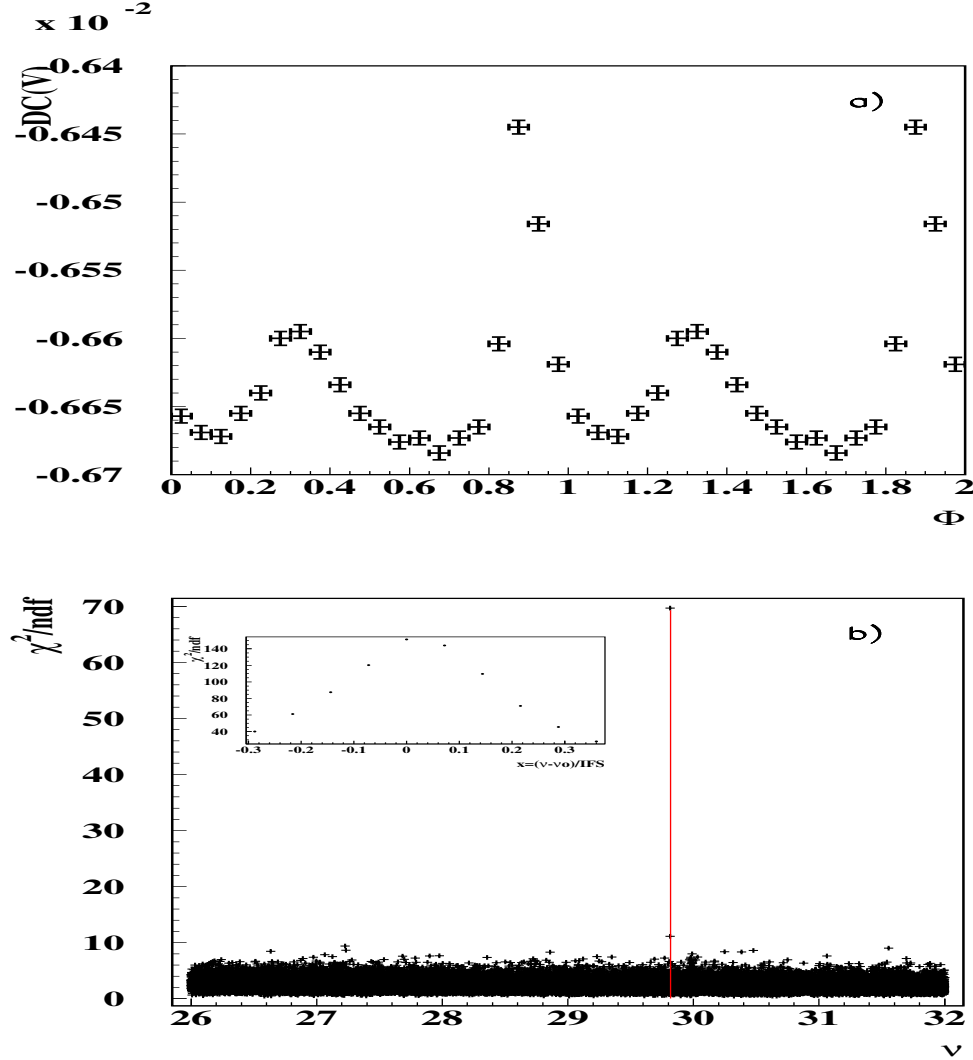


Fig. 1. Observations of the optical Crab pulsar with CT1. On the top, the two peaks are clearly visible in the light curve with a separation in phase of ~ 0.4 . On the bottom (b) the excess in the reduced χ^2 searching in a wide range of frequencies, using the 15th October ephemeris. The difference between the frequency found with 15th October ephemeris and 15th November is consistent with the drift in the digitization card clock. The inset shows the value of the reduced χ^2 vs. $x = (\nu - \nu_o)/IFS$, where ν_o corresponds to 15th October.

5 Determination of the galactic anticenter LONS

In this section, we will determine the galactic night sky background around the Crab pulsar by using the pulsed fraction of the signal detected in CT1. This result will be checked by comparing it with the LONS calculated from the intensity of the CT1 camera response to ζ Tauri.

It is possible to derive the pulsed fraction of the optical Crab signal from the density function of the light curve. This density function results from the unpulsed background (nebula and night sky background) with fraction $1 - p$ and the density function $f_s(\theta)$ of the pulsed signal coming from the Crab pulsar:

$$f(\theta) = pf_s(\theta) + (1 - p), \quad (5)$$

where p is the pulsed fraction, if the density function is normalised to unity. The off-pulse interval (region of minimum intensity) following the interpulse and preceding the main pulse was used to estimate the pulsed fraction.

Taking the proper errors in each bin into account (exploiting the Central Limit Theorem), we derive a pulsed fraction for the Crab of $p = (5.00 \pm 0.02) \times 10^{-4}$.

To check the stability of the signal with time, the total data set of two hours was divided into six data sets of 20 minutes each. The pulsed fraction, as well as the ratio between the first and second peaks was calculated for each subset. The results are shown in the Figure 2. It is clear that both parameters appear to be stable within statistical errors. An indirect indication of the stability of telescope tracking during two hours observations is shown by the constant value of the pulsed fraction.

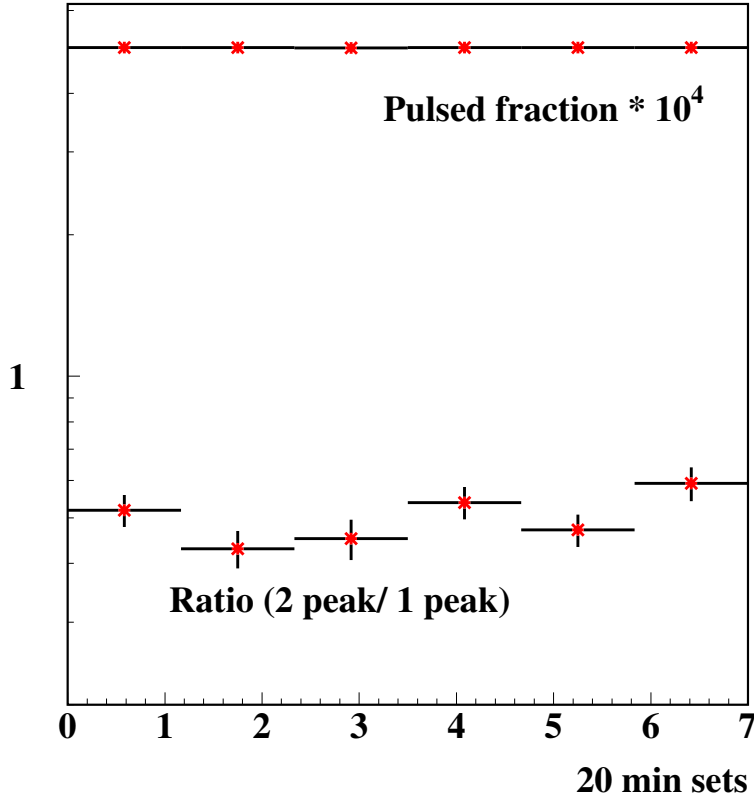


Fig. 2. *Test of the telescope response to the pulsed signal in two hours. The pulsed fraction times 10^4 for the set of 20 min observation slices is at the top of the plot. The ratio between the first and second peak remains constant with time, as shown at the bottom.*

There is a difference between the measured CT1 Crab pulsed fraction of 5×10^{-4} and the theoretical value given in the previous section ($p_{theor} = 8.6 \times 10^{-4}$). The discrepancy can be attributed to the uncertainty on the measurement of the LONS, the influence of the mirror point spread function (PSF), and the fact that the Crab pulsar was off-center due to uncorrected but constant mispointing in the telescope.

To determine the galactic LONS around the Crab pulsar, we first need to know the PSF of the telescope. With this purpose, we used the bright source

ζ Tauri to measure the amount of light, which falls within the central pixel.

The same source is used to measure the telescope misspointing.

The selected source was ζ Tauri (ICRS 2000.0 coordinates are $5^h:37':38.70''$ RA, $21^\circ:08':33.0''$ DEC with a magnitude of 2.84), which also appears in the camera field of view when CT1 points to the Crab pulsar. The advantages of using this reference star are the following: (1) we know its photometric properties and (2) the atmospheric extinction will be the same for both the pulsar and the star, given their proximity in the sky. The PSF of the telescope was determined from a raster scan with a resolution finer than the 0.25° pixel size (see Fig. 3(a)), with zero position corresponding to the true position of the star). Since the telescope is focused to ~ 6 km (maximum $\hat{C}$ erenkov shower), the image of a point-like source will be smeared out around 0.14° .

The signal from ζ Tauri is slightly shifted to lower right ascension values, due to the telescope mispointing. This was accounted for by calculating the maximum value of the signal, and measuring the difference between the shifted measurement and the true position. A schematic figure of the CT1 central pixel is also drawn in Figure 3. To calculate the true signal collected by the central pixel, the signal intensity was fitted to a bidimensional Gaussian (fig. 3 (b, c)), and the fraction of the signal within the central pixel was measured to be:

$$\Pi = \frac{\int_0^{r_{pixel}} \int_0^{2\pi} G(r, \theta) dr d\theta}{\int_0^\infty \int_0^{2\pi} G(r, \theta) dr d\theta} = 69\%. \quad (6)$$

From the misspointing correction factor Π , it was possible to derive the LONS rate from:

$$p = \frac{R_{pulsed}}{R_{pulsed} + R_{nebula} + R_{LONS}}, \quad (7)$$

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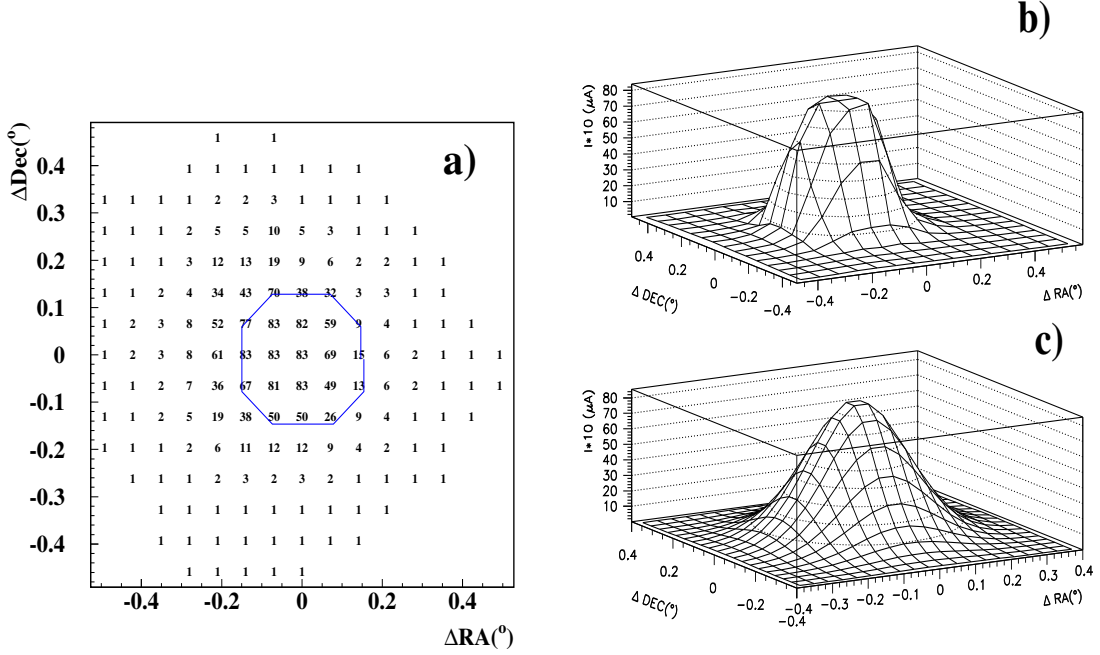


Fig. 3. DC currents due to ζ Tauri on the CT1 central pixel, in $0.1 \mu\text{A}$ units. Left (a): ζ Tauri signal in the CT1 camera, with background already subtracted. The zero position corresponds to the true position of the star in equatorial coordinates (maximum intensity for zero mispointing). The approximate shape of the central pixel is drawn to illustrate the size of the PSF relative to the pixel size. Top right: (b) ζ Tauri observed intensity in a 2-dimensional plot. Bottom right: (c) The same signal, fitted with a bidimensional Gaussian.

where p is the empirical pulsed fraction, R_{pulsed} is the expected rate (table 2) of Crab pulsed events, $R_{nebulas}$ the expected rate from the Crab nebula (table 2) and R_{LONS} is the galactic background rate around the Crab nebula.

Finally, assuming that the contribution of the pulsed signal is negligible compared to the background due to the nebula and the LONS we have:

$$p = \frac{\int F_{\nu}^{pulsed} \cdot \eta \cdot Q \cdot A_{ct1} \epsilon_{\gamma v} \Pi d\nu}{\int F_{\nu}^{nebulas} \cdot \eta \cdot Q \cdot A_{ct1} \epsilon_{\gamma v} \Pi d\nu + \int F_{\nu}^{LONS} \cdot \Omega \cdot \eta \cdot Q \cdot A_{ct1} \epsilon_{\gamma v} d\nu}, \quad (8)$$

where η is the mirror reflectivity, Q the quantum efficiency of the PMs, A_{ct1} is

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the CT1 mirror area, Π is the fraction of light from a point source captured by the central pixel, Ω is the solid angle subtended by the central pixel, and $\epsilon_{\gamma v}$ is a conversion factor between number of photoelectrons and voltage measured by the digitization card, which cancels out in the above expression.

Therefore, using the pulsed fraction of the total Crab signal, the frequency integrated LONS (weighted with the frequency dependency of the PM quantum efficiency) is:

$$I_{LONS} = \int F_{\nu}^{LONS} d\nu = (4.6 \pm 0.1) \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \cdot \text{s}^{-1} \cdot \text{sr}^{-1} \quad (9)$$

To cross-check this result, we used the signal detected by CT1 from ζ Tauri in the central pixel (fig. 3): If i_{cp} is the total observed current intensity from the direction of ζ Tauri in the central pixel, $i_{\zeta \text{ Tauri}}$ is the expected contribution to the central pixel current from ζ Tauri, and $\epsilon_{\gamma i}$ is the conversion factor from photoelectrons to current (in μA), which is calculated from

$$\epsilon_{\gamma i} \Pi \int F_{\nu}^{\zeta \text{ Tauri}} \cdot \eta \cdot Q \cdot A_{ct1} d\nu = i_{\zeta \text{ Tauri}} = i_{cp} - i_{LONS} \sim i_{cp}. \quad (10)$$

The last approximation is valid since ζ Tauri is a very bright star rendering the contribution from the LONS in the central pixel negligible. $F_{\nu}^{\zeta \text{ Tauri}}$ is tabulated including atmospheric extinction. We obtain a measured value of $\epsilon_{\gamma i} = 2.95 \times 10^{-9} \text{ mA} \cdot \text{s} \cdot \text{phe}^{-1}$. From this value we obtain the LONS by subtracting the PSF of ζ Tauri from a square degree image centered on this star (shown in Figure 3). The error on the LONS is minimal near the outer wings of the PSF, since the LONS is dominant in that area. An annular region is used to estimate the LONS, giving

$$I_{LONS} = \int F_{\nu}^{LONS} d\nu \sim (4.3 \pm 0.1) \times 10^{12} \text{ ph} \cdot \text{m}^{-2} \cdot \text{s}^{-1} \cdot \text{sr}^{-1}, \quad (11)$$

after normalising to the solid angle subtended by the selected annular region.

The two values for the galactic I_{LONS} (eq. 9 & 11) calculated by two independent methods agree within $\sim 7\%$ and are ~ 2.7 times higher than the previous measurement (7), used for our theoretical predictions as we expected. The LONS intensity calculated from the pulsed fraction is the most accurate, being the correction factor Π uncertainty the highest, since the LONS determination from ζ Tauri also includes the uncertainty in the background region around the star.

6 Sensitivity for Pulsed Detection in Optical

The sensitivity of CT1 telescope to detect the Crab optical pulse can be expressed by the "standard deviation" s of the signal relative to the behavior of the reduced χ^2 at frequencies different from the Crab frequency:

$$s = \frac{\chi_{red}^2 - \mu_{background}}{\sigma_{background}} \quad (12)$$

for each observation time interval, where $\mu_{background}$ is the mean value of χ_{red}^2 at such offset frequencies and $\sigma_{background}$ the corresponding standard deviation, which is constant for all selected times T . The total 2 hours observation time was split into $L = 7200 \text{ s}/T$ independent datasets, each of length T , from which the average value of the statistic s was calculated for $T = 50 \text{ s}$ to 7200 s in steps of 50 s . It can be shown that this average value of s is:

$$s \sim p^2 \cdot T \cdot g(k, f_s), \quad (13)$$

where k is the number of bins in the phaseogram, f_s the signal pulse shape and $g(k, f_s)$ a function depending on these quantities. Note that s is not a Gaussian standard deviation, since the number k is too small, but we expect that $\langle s \rangle \propto \sigma^2$, where $\sigma \sim p \cdot \sqrt{T}$ would be the Gaussian significance of a DC detection of the same signal. Figure 4 shows a graph for $\log \langle s \rangle$ vs. $\log T$, showing a linear relationship between s and T as predicted.

Using the calculated value of the galactic LONS from the pulse analysis, we estimate the required time for MAGIC to detect the Crab pulsar in the optical, given the much larger mirror area, improved quantum efficiency, and pixel size (0.1°). By scaling the graph of $\log \langle s \rangle$ vs. $\log T$ accordingly, Figure 4 shows that MAGIC can detect the Crab pulsar in 30 s at the 5σ level. The line on top shows the prediction for a PM without inherent fluctuations (eq. 3), using the galactic LONS calculated.

7 Conclusions

The Crab pulsar was measured in the optical with the HEGRA CT1 telescope (the smallest Čerenkov telescope making this detection). A 5σ detection requires an observation time of 200 seconds.

Since the design of the CT1 PM base is very similar to that of MAGIC, we extrapolate CT1 results and get a 5σ detection with MAGIC (~ 25 times larger dish area than CT1) in a substantially shorter time of 30 sec. Considering that it is possible to reduce the LONS using an array of HPD's, combined with the superior PSF of MAGIC, it could still be possible to resolve single pulses.

The use of a PM base with the technical modifications described above is

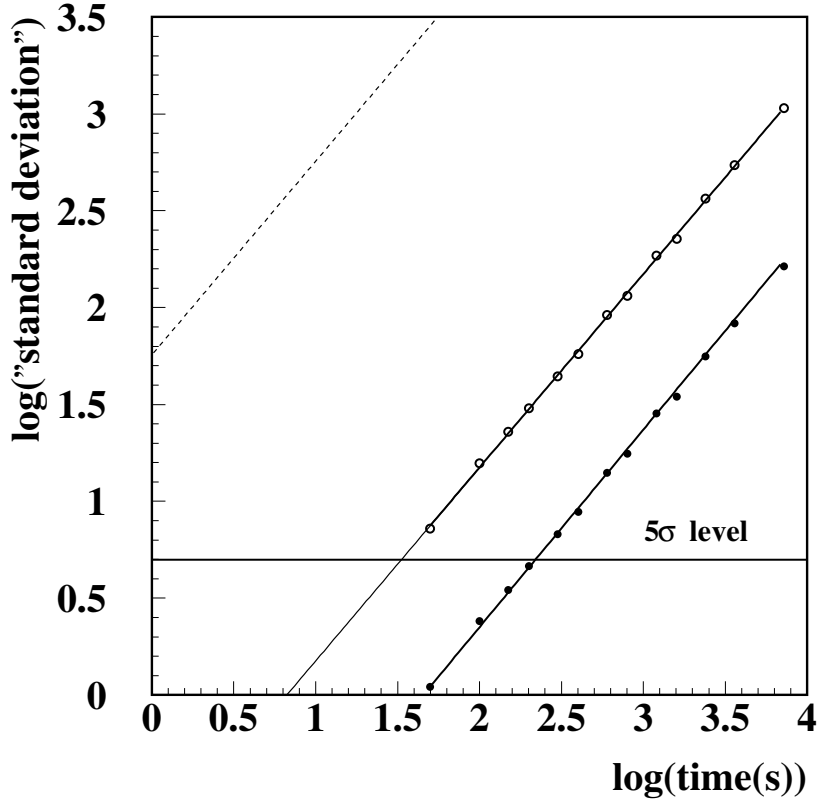


Fig. 4. *log of "standard deviation" s of the signal related to the background versus observation time. The CT1 empirical data are represented with filled bullets while the empty ones represent the extrapolation for the MAGIC telescope. The horizontal line shows the "5 σ " level.*

incompatible with normal VHE observations. The modification could only be considered for the central pixel of the telescope camera.

The galactic LONS background on the surroundings of the Crab region was calculated using the Crab pulsed fraction. Only 7% disagreement was found with a second method using a bright reference star and a different background region (still within the camera field-of-view). Our results are however on average 2.7 times higher than the extragalactic LONS from (7). The advantage of this new method is that it relies on the stability of the clock, and is indepen-

dent of other sky-techniques in the DC mode, provided that the PSF is known.

The determination of the LONS for the Crab region is important, since the threshold energy determination and the spectrum depend on the level of the LONS.

Simultaneous Crab pulsar optical data and Čerenkov VHE observations could be very important for monitoring atmospheric extinction. In fact, high quality observations should show a very specific zenith angle dependency of the pulsed fraction. Thus, the "seeing condition" at a given time may also be derived from the observed pulsed fraction corrected for atmospheric extinction. Furthermore, apart from using optical observations to provide a timing ephemeris for VHE observations.

We can also monitor deviations from the normal cubic law for pulsar spindown (assuming the pulsar braking index is detectable) as a result of timing noise and glitches. Thus, frequent monitoring of the Crab by MAGIC (for calibration purposes) should also improve the chance of detecting glitches. In fact, very specific software can be written to test glitching behavior during observations. A specially designed central pixel detector with high quantum efficiency, and relatively low noise will be very useful for monitoring the state of optical emission from γ -ray sources such as cataclysmic variables, X-ray binaries, microquasars, AGNs and possibly GRBs.

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