



060043851R

North-West University
Mafikeng Campus Library

MULTI-COLOUR PHOTOMETRY OF β CEPHEI STARS IN THE LARGE MAGELLANIC CLOUD

By

Malefetsane Daniel Nhlapo

Student No: 22894683

A dissertation submitted in fulfillment of the requirements for the
M.Sc. degree in the Faculty of Agriculture, Science & Technology at
NORTH WEST UNIVERSITY

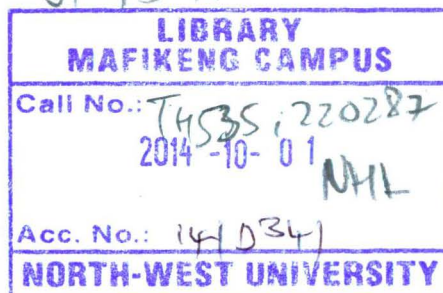
November 2011

Supervisors:

Prof. Medupe R.¹

Dr. Balona L.A.²

619591081



¹Physics Department, North West University

²South African Astronomical Observatory

Abstract

The β Cephei stars are a sub-group of pulsating B-stars that pulsate with periods of between 1.6 hours and 7.7 hours. These population I stars have spectral types between O9 and B3 and the luminosity classes ranging from I to V. The pulsations in β Cephei stars are driven by the κ -mechanism, which operates at the temperature of 2×10^5 K. This mechanism in β Cephei stars is highly sensitive to the metal content of the star. Theoretical models showed that the κ mechanism in β Cephei stars depends strongly on metallicity, and that β Cephei pulsations disappear at metallicities of less than 0.01. However, in 2002 β Cephei stars were discovered in the Large Magellanic Cloud (LMC). This discovery created a discrepancy between theory and observations because the metallicity of the LMC was measured to be less than 0.01.

The initial aims of this thesis were to search for new β Cephei stars in the LMC using the OGLE database, to observe the target stars and to perform period analyses of the discovered extragalactic β Cephei stars. However, due to time constraints we only present colour-magnitude diagrams and M_V vs (B-V) diagrams of observed LMC stars. To obtain these diagrams, we searched for candidate β Cephei stars in the LMC. The search was done by performing Fourier analysis of the light curves we obtained from OGLE II database. From the list of promising β Cephei candidates, we selected those with the largest pulsation amplitudes. The observations of the three selected candidates were done in three observation campaigns. The data obtained was reduced using three CCD reduction programs. The first program (Duphot) was found to be not suitable for dealing with our crowded field images. The data was eventually reduced using DoPhot and ISIS. Since ISIS was designed specifically for crowded field images, comparing its output to the output from DoPhot showed that DoPhot was capable of producing reasonable photometry. However, the reduced magnitudes from both programs could not be compared directly due to different magnitude units produced.

The reduced output from DoPhot was used to plot the colour-magnitude diagrams of LMC fields we observed. Since we did not observe standard stars, we used the

magnitudes of stars from the OGLE II fields that correspond to the stars in the fields we observed, in order to standardize magnitudes. The resulting standardized magnitudes were used to plot colour-magnitude diagrams and the M_V vs $(B-V)$ diagrams of the fields we observed. By marking the M_V and $(B-V)$ ranges of β Cephei stars, we found that there are eleven stars from LMC_T1 and seven stars from LMC_T13 located within these ranges. This gives a preliminary indication that we have found 18 β Cephei stars in LMC. Conclusive results will be made once we perform period analysis of the data of the 18 β Cephei candidate stars.

Acknowledgements

- I am grateful to my supervisor, **Prof. R. Medupe** for suggesting the topic of this research and also for guiding and inspiring me through this project.
- I am very grateful to **Dr. L.A. Balona** for his major and important contribution to this project. The programs he provided for reducing and analysing the data were of great importance. All the discussions and the guidance he gave me are highly appreciated.
- I also thank **Mr. Francois van Wyk** for his unlimited kindness and help, observations for this project would have never been so easy and enjoyable without his assistance.
- I also thank **Department of Physics members at Mafikeng campus of North West University** for their everlasting help and understanding whenever I wanted to print and scan documents. I particularly want to thank **Mr. Makgathe** for his encouraging words, his morning visits to our research room were inspiring.
- I would like to express my humble thanks to **National Astrophysics and Space Science Programme** for the financial support. I also thank my **fellow students** for all their help and fruitful discussions, their presence around me has always been my driving mechanism.

Dedications

This work would not have been possible without the support of the following:

- **Mmē Mamatshediso Letia Nhlapo**, my mother, for your unlimited love and advice. Ke a leboha ka kgodiso e kgabane eo o nnehileng yona, lerato le tshireletso tsa hao ke tsona tse nkentseng seo ke leng sona kajeno, o Mmē wa mmakgonthe.
- **Ntate Peter Albert Nhlapo**, my late father. Ka hara dintho tsohle tsoe o nketseditseng tsona ha o sa phela, ke leboha thuto eo o nthutileng yona: hlomphe, no son could ever ask for a better father, Ke a leboha.
- **Bokgaitsemi, ausi Puseletso Nhlapo, Mathapelo Nhlapo le Mamello Nhlapo**, my sisters, boteng ba lona bophelong ba ka bo entse hore ke be le sepheo sa ho phela, and I trust that this work will inspire you to be strong and always aim for the stars in everything you are doing in your lives. Ke a leboha bana beso.
- **Batjhana ba ka, Lehlohonolo Nhlapo, Monaheng Mokoena, Mohlomi Mokoena, Mohlominyana Mokoena le Katleho Nhlapo**, my nephews, boteng ba lona bo etsa hore ke be malome ya hlomphehang, thinking about you has always reminded me that there are people looking up to me.
- **Kedibone Ratsaana**, my best friend, your presence around me has been the source of my happiness and strength, no man could ever ask for a better friend, Tsala ke lebogela tshegetso ya gago. Tiro e e ne seke ya kgonagala ntleng le tlhokomelo le lorato lwa gago.

And to all my extended families and friends, this work is for all of you too, hopefully to encourage you to go further, and reach for your dreams. **Kgotso, Pula, Nala.**

Contents

1	Introduction	2
1.1	B stars in general	5
1.2	β Cephei stars	6
1.3	Photometric properties of β Cephei stars	7
1.3.1	β Cephei stars in the Milky Way	8
1.3.2	Summary of photometric properties of Galactic β Cephei stars .	20
1.3.3	β Cephei stars in Magellanic Clouds	23
1.4	Spectroscopic properties of β Cephei stars	29
1.5	Slowly Pulsating B (SPB) stars	38
1.6	Pulsating Be stars	40
1.7	The mechanism that drives pulsations in β Cephei stars	42
1.8	A description of the OGLE database	47
2	Target Selection and Observations	50
2.1	Target Selection	50
2.2	Observations	74
2.2.1	The November 2010 campaign	74
2.2.2	Observing Procedure	78
2.2.3	The February 2011 campaign	85
2.2.4	The March 2011 campaign	86
3	Data Reductions	90
3.1	DUPHOT	90

3.2	DoPhot	96
3.2.1	Point Spread Function (PSF) fitting Photometry	96
3.2.2	Aperture Photometry	99
3.2.3	Preliminary steps	102
3.2.4	Extracting the photometry	103
3.3	ISIS	105
3.3.1	Image Registration	105
3.3.2	Image Subtraction	106
3.3.3	Photometry	107
4	Analysis and Results	109
4.1	Comparison of DoPhot and ISIS outputs	109
4.2	Colour-magnitude diagrams	119
4.3	Standardizing B,V and I magnitudes	127
5	Conclusion	146
5.1	Future work	147
A		152
A.1	Target selection	152
A.2	Analysis and Results	154
A.2.1	Comparison of DoPhot and ISIS light curves	154
A.2.2	“cat“ files manipulations	162
A.2.3	Standardization of magnitudes	163
A.3	Observations	178
A.3.1	Visibility Curves Of Observed Targets	178
A.4	Data Reductions	191
A.4.1	Contents of dophot.inp	191
A.4.2	The header of the “sum” files	193
A.4.3	The contents of “ccdpick.out” file	195
A.5	The 1.0-m telescope	196

A.5.1	Setting the zero points	202
-------	-----------------------------------	-----

List of Figures

- 1.1 A figure showing instability strips of various pulsating stars in the HR diagram. The instability strip of β Cephei stars is on the upper main-sequence. The dashed line shows the zero-age main sequence, the continuous curves are selected evolution tracks, the dot-dashed line is the horizontal branch and the dotted curve is the white-dwarf cooling curve (taken from Christensen-Dalsgaard 1999). 3
- 1.2 A plot showing light variations of HD 21803. The vertical axis has the magnitude differences of HD 21803 and comparison star in yellow light (taken from Lynds 1959). 10
- 1.3 A Stromgren b-filter light curve of HR 6684 measured by Jerzykiewicz (1972), during a search for new β Cephei stars. The bottom dotted-curve is the magnitude difference of comparison stars (Taken from Jerzykiewicz 1972). . 12
- 1.4 A plot showing HR-diagram of stars Balona (1977) observed. The curved line is a semi-empirical Zero-Age Main Sequence (ZAMS), dashed line next to ZAMS is a theoretical ZAMS. The upper two lines show theoretical region of post main sequence evolution (Taken from Balona 1977). 13
- 1.5 A plot showing frequency spectra of HD101794. The panels from the top show the spectra in the sequence of prewhitening, the top panel marked (a) shows the first frequency, middle panel (b) shows the second frequency with the first frequency peak removed and the last bottom panel (c) shows a spectrum with all significant peaks removed (taken from Drobek et al. 2010). 15

1.6	A plot showing frequency spectra of HD101838 in the sequence of prewhitening from the top spectrum, top panel (a) shows the only frequency peak found and the bottom panel (b) is the spectrum with the peak removed and mostly showing noise (taken from Drobek et al. 2010).	16
1.7	A plot showing light curve (left) and frequency spectra (right) of β Cephei star NGC 637 4. The frequency spectra are in the order of prewhitening from top to bottom (taken from Handler & Meingast 2011).	18
1.8	A plot showing light curve (left) and frequency spectra (right) of β Cephei star NGC 637 1. The frequency spectra are in the order of prewhitening from top to bottom (taken from Handler & Meingast 2011).	18
1.9	A plot showing light curve (left) and frequency spectra (right) of β Cephei star NGC 637 7. The frequency spectra are in the order of prewhitening from top to bottom (taken from Handler & Meingast 2011).	19
1.10	A plot showing light curve (left) and frequency spectra (right) of β Cephei star NGC 637 138. The frequency spectra are in the order of prewhitening from top to bottom (taken from Handler & Meingast 2011).	19
1.11	A plot showing light curve of suspected β Cephei star NGC 637 6 (taken from Handler & Meingast 2011).	20
1.12	A 3D plot showing the distribution of β Cephei stars according to their spectral class and luminosity class (taken from Stankov & Handler 2005).	21
1.13	A plot showing the distribution of Galactic β Cephei stars with pulsation period by 2005 (taken from Stankov & Handler 2005).	22
1.14	A plot showing the distribution of Galactic β Cephei stars in V magnitudes by 2005 (taken from Stankov & Handler 2005).	22
1.15	A plot showing colour-magnitude relation of the sample of stars from the LMC. The three discovered extragalactic β Cephei stars are shown as filled circles, empty circles are suspected β Cephei stars (Taken from Pigulski & Kolaczowski 2002).	25

1.16	A plot showing colour-magnitude diagram of the sample of LMC stars, from which 64 β Cephei pulsators were discovered. These 64 β Cephei stars are represented by circles and suspected β Cephei stars are represented by squares (taken from Pigulski & Kolaczowski 2004).	27
1.17	A plot showing colour-magnitude diagrams (i.e. (B-V) on the left panel and (V-I) on the right panel) of a sample of LMC and SMC stars. The discovered β Cephei stars are shown with open circles and dots in squares for LMC and SMC, respectively. These diagrams also show discovered Be stars (encircled dots) and the vertical line on the right panel separates Be stars and non-Be stars (taken from Kolaczowski et al. 2006).	28
1.18	A figure showing the <i>provisional</i> velocity-curve for β Cephei discovered by Frost (1902). The circles are the data points and the curve is the fit to the data points. One should note a mistake Frost (1906) did by expressing velocity in units of a distance in vertical axis.	30
1.19	A figure showing a continuous curve which represents the radial velocities of a pulsating star from SiIII, OII, CII. The dashed curve is the radial velocity curve for HI (taken from Struve 1955).	33
1.20	A figure showing a continuous curve representing the radial velocity of SiIII, OII, CII and a dashed curve representing the radial velocity of HeI (taken from Struve 1955).	34
1.21	A figure showing theoretical instability strips and observational data of β Cephei stars in the $\log g$ vs $\log(T_{\text{eff}})$ diagram. The left panel is for different metallicities while the right panel is for different clusters (taken from Daszyńska-Daszkiewicz & Niemczura 2003).	36
1.22	Figures showing HeI 6678 Å line profile variations of V986 Oph (left) and of KP Per (right) (taken from Stateva et al. 2010).	37
1.23	A figure showing consecutive steps of prewhitened Fourier periodograms of the multi-periodic SPB star, SC1-223227 in the LMC (taken from Kolaczowski et al. 2004).	39

1.24	A figure showing the HR diagram of β Cephei stars (dots) and theoretical instability strips of the radial fundamental mode. The diagonal straight line represents the Zero Age Main Sequence (ZAMS) for a metallicity of 0.03, the solid line curve is the theoretical instability strip for a metallicity of 0.03 and the dashed curve is the theoretical instability strip for a metallicity of 0.04 (taken from Moskalik & Dziembowski 1992).	45
1.25	The figure showing the LMC and twenty-seven rectangular regions within which OGLE observations were made (see http://ogledb.astrouw.edu.pl/ogle/photdb/).	49
2.1	A figure showing the OGLE Photometry Query Page where the parameters on Table. 2.1 were entered to extract the data for target selection (see http://ogledb.astrouw.edu.pl/ogle/photdb/).	52
2.2	The top plot shows the OGLE light curve of <code>lmc_sc4_i_296366</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 9.278 c/d. The highest peaks between 5 and 10 c/d look promising. Also notice a noise level which is less than 2.5 mmag on the spectrum.	55
2.3	The top plot shows the OGLE light curve of <code>lmc_sc11_i_284846</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 8.224 c/d. The highest peak just around 5 c/d looks promising and the other two promising peaks are between 0 and 5 c/d, even though they might be aliases given the noticeable gaps on the light curve.	56
2.4	The top plot shows the OGLE light curve of <code>lmc_sc13_i_270406</code> , and the bottom plot is its frequency spectrum. The highest peak between 5 and 10 c/d looks promising.	57
2.5	The top plot shows the OGLE light curve of <code>lmc_sc13_i_218743</code> , and the bottom plot is its frequency spectrum. The maximum peak between 10 and 15 c/d is promising. The other noise-outstanding peaks could be aliases as it can be seen from the light curve that there are gaps.	58

2.6	The top plot shows the OGLE light curve of <code>lmc_sc10_i_212836</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 7.880 c/d. There are two peaks which are promising on this spectrum, the first is between 5 and 10 c/d and the second is between 0 and 5 c/d. The latter is however as high as the noise level.	59
2.7	The top plot shows the OGLE light curve of <code>lmc_sc7_i_207004</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 9.338 c/d. The maximum peak around 5 c/d is promising and also notice the noise level which is around 2 mmag on this spectrum. The other peaks might be aliases given the gaps on the light curve.	60
2.8	The top plot shows the OGLE light curve of <code>lmc_sc11_i_38067</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 8.224 c/d. The highest peak between 5 and 10 c/d is promising.	61
2.9	The top plot shows the OGLE light curve of <code>lmc_sc13_i_53816</code> , and the bottom plot is its frequency spectrum. There are three promising peaks between 0 and 5 c/d and also just above 5 c/d.	62
2.10	The top plot shows the OGLE light curve of <code>lmc_sc3_i_108137</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 9.40 c/d. The highest peak between 0 and 5 c/d is promising and other peaks on both sides of the highest peak look more like aliases. Also notice the noise level on this spectrum which is around 5 mmag.	63
2.11	The top plot shows the OGLE light curve of <code>lmc_sc4_i_113267</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 9.278 c/d. The highest peak between 0 and 5 c/d looks promising and the spectrum has noise level of about 3 mmag.	64
2.12	The top plot shows the OGLE light curve of <code>lmc_sc10_i_132705</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 7.880 c/d. The highest peak between 5 and 10 c/d looks promising. The other peaks are hidden by the noise level of about 3 mmag on this spectrum.	65

2.13	The top plot shows the OGLE light curve of <code>lmc_sc9_i_140826</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 6.846 c/d. The peak between 5 and 10 c/d looks promising.	66
2.14	The top plot shows the OGLE light curve of <code>lmc_sc13_i_142737</code> , and the bottom plot is its frequency spectrum. The highest peak between 5 and 10 c/d looks promising. The other peaks look more like aliases and are as high as the noise level.	67
2.15	The top plot shows the OGLE light curve of <code>lmc_sc1_i_179848</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 6.369 c/d. The highest peak at around 5 c/d looks promising. The other peaks on the right of the highest peak look like aliases given the noticeable gaps on the light curve.	68
2.16	The top plot shows the OGLE light curve of <code>lmc_sc13_i_194180</code> , and the bottom plot is its frequency spectrum. The highest peak between 5 and 10 c/d looks promising. The other highest peaks might be aliases and the other noticable feature is the noise level of about 3.3 mmag.	69
2.17	The top plot shows the OGLE light curve of <code>lmc_sc7_i_199968</code> , and the bottom plot is its frequency spectrum with a Nyquist frequency of 9.338 c/d. The highest peak between 0 and 5 c/d looks promising. The other peaks look like aliases. Also notice the noise level of about 2.8 mmag on the spectrum.	70
2.18	A figure showing the OGLE finding charts of all six β Cephei candidates selected for observations, from top left, clockwise is the image of T1, T4, T9, T13, T11, and T6. Each candidate is indicated by an arrow on the images (see http://ogledb.astrouw.edu.pl/ogle/photdb/).	72
2.19	Figures showing the flat field frames from four Johnson filters (UBVI) obtained during the first observation run (05 November 2010). In clockwise direction, U-flat is the top-left frame, B-flat is top-right, I-flat is bottom-right and V-flat is bottom-left. The black spots on the frames are bad pixels which should be noted during observation, so that the target is not placed on any of those spots. The scale below each flat frame is the brightness scale and it increases from left to right.	80

2.20	A figure showing the CCD frames of T1 (top) and T13 (bottom) from the LMC, observed during November 2010 observation run. T13 is located in the cluster (pointed by an arrow) seen on its frame and it can be seen that both frames are very crowded and T1 is the star pointed by an arrow towards the left of the frame.	82
2.21	A figure showing the finding charts of the stars close to T1 (top) and T13 (bottom) that were used to locate our targets (see http://simbak.cfa.harvard.edu/simbad/). 83	
2.22	A figure showing the frame taken for T6, which is located towards the bottom-center of the frame and indicated by an arrow.	88
3.1	A figure showing the ds9 interface and how it displays the FITS images (see: http://hea-www.harvard.edu/RD/ds9/user/gui/index.html#components). .	92
3.2	A figure showing the light curves of first four brightest stars from the results of November 2010 observation campaign. On the right panel are the raw light curves and on the left panel are light curves after differential photometry has been applied, and star 1 was used as a comparison to demonstrate an effect of differential photometry on light curves of other three stars as targets.	95
3.3	Plots of intensity as a function of position in CCD (in pixels) from above the Earth's atmosphere (left panel) and from the Earth's surface (right panel). These plots represent atmospheric and instrumental effects on the light from a point source star.	97
3.4	A figure showing two CCD images of the LMC-T6 field, the image on left panel is the raw image and the right panel image is flat-fielded and bias removed. Bias regions are the black regions on left and right edges of the raw image. .	102
4.1	A figure showing the light curves of star 8 from DoPhot reduction outputs in different nights.	111
4.2	A figure showing the light curves of star 1 from ISIS reduction outputs, this is the same star as in Fig.4.1.	111

4.3	A figure showing the light curves of star 1885 from DoPhot reduction outputs in different nights.	112
4.4	A figure showing the light curves of star 2 from ISIS reduction outputs, this is the same star as in Fig.4.3.	112
4.5	A figure showing the light curves of star 343 from DoPhot reduction outputs in different nights.	113
4.6	A figure showing the light curves of star 4 from ISIS reduction outputs, this is the same star as in Fig.4.5.	113
4.7	A figure showing the light curves of star 7 from DoPhot reduction outputs in different nights.	114
4.8	A figure showing the light curves of star 6 from ISIS reduction outputs, this is the same star as in Fig.4.7.	114
4.9	A figure showing the light curves of star 6844 from DoPhot reduction outputs in different nights.	115
4.10	A figure showing the light curves of star 274 from ISIS reduction outputs, this is the same star as in Fig.4.9.	115
4.11	A figure showing the light curves of star 6942 from DoPhot reduction outputs in different nights.	116
4.12	A figure showing the light curves of star 844 from ISIS reduction outputs, this is the same star as in Fig.4.11.	116
4.13	A figure showing the light curves of star 6324 from DoPhot reduction outputs in different nights.	117
4.14	A figure showing the light curves of star 1113 from ISIS reduction outputs, this is the same star as in Fig.4.13.	117
4.15	A figure showing the light curves of star 6848 from DoPhot reduction outputs in different nights.	118
4.16	A figure showing the light curves of star 1842 from ISIS reduction outputs, this is the same star as in Fig.4.15.	118
4.17	A plot showing the relation between V-magnitude and B-V-colour of the LMC-T1 field. The plotted data was reduced using DoPhot.	120

4.18	A plot showing the relation between V-magnitude and U-I-colour of the LMC_T1 field. The plotted data was reduced using DoPhot.	121
4.19	A plot showing the relation between V-magnitude and V-I-colour of the LMC_T1 field. The plotted data was reduced using DoPhot.	121
4.20	A plot showing the relation between V-magnitude and U-B-colour of the LMC_T1 field. The plotted data was reduced using DoPhot.	122
4.21	A plot showing the relation between V-magnitude and B-V-colour of the LMC_T13 field. The plotted data was reduced using DoPhot.	123
4.22	A plot showing the relation between V-magnitude and U-B-colour of the LMC_T13 field. The plotted data was reduced using DoPhot.	123
4.23	A plot showing the relation between V-magnitude and U-I-colour of the LMC_T13 field. The plotted data was reduced using DoPhot.	124
4.24	A plot showing the relation between V-magnitude and V-I-colour of the LMC_T13 field. The plotted data was reduced using DoPhot.	124
4.25	A plot showing the effect of flat-fielding raw images with a normalized flat field frame of all flat field frames of an entire observation campaign combined (right image), and raw image flat-fielded with normalized flat field frames of a week (left image).	126
4.26	Comparison of star patterns in a CCD images and the same pattern in the OGLE maps. The diagrams represent the stars and distances presented in equation (4.3). The left panel diagram represents any stars A, B and an arbitrary reference star R on a the CCD image and right panel diagram represents corresponding stars on the OGLE map.	128
4.27	The top image shows stars in the CCD frame we observed in the LMC_T1 field. The bottom image shows the same group of stars in the OGLE II map identified for calculating coefficients in equations (4.1) and (4.2). Identified stars are pointed with arrows and marked with letters.	130

4.28	The top image shows stars in the CCD frame we observed in the LMC_T13 field. The bottom image shows the same group of stars in the OGLE II map identified for calculating coefficients in equations (4.1) and (4.2). Identified stars are pointed with arrows and marked with letters.	131
4.29	A figure showing colour-magnitude diagrams of LMC_T1 (upper panels) and the LMC_T13 (lower panels) from standardized magnitudes. It can be seen that these diagrams are consistent with those plotted with unstandardized magnitudes on Figs. 4.17, 4.19, 4.21 and 4.24.	140
4.30	Figures showing absolute magnitude (M_V) against the colour (B-V) for the LMC_T1 (top panel) and the LMC_T13 (bottom panel) fields. Horizontal lines mark M_V range of β Cephei stars. Suspected β Cephei stars are marked with crosses inside circles.	143
A.1	Plots showing altitude as a function of time, the solid curve marks the altitude of T1 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	179
A.2	Plots showing altitude as a function of time, the solid curve marks the altitude of T1 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	180
A.3	Plots showing altitude as a function of time, the solid curve marks the altitude of T1 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	181
A.4	Plots showing altitude as a function of time, the solid curve marks the altitude of T1 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	182
A.5	Plots showing altitude as a function of time, the solid curve marks the altitude of T13 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	183
A.6	Plots showing altitude as a function of time, the solid curve marks the altitude of T13 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	184

A.7	Plots showing altitude as a function of time, the solid curve marks the altitude of T13 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	185
A.8	Plots showing altitude as a function of time, the solid curve marks the altitude of T13 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	186
A.9	Plots showing altitude as a function of time, the solid curve marks the altitude of T6 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	187
A.10	Plots showing altitude as a function of time, the solid curve marks the altitude of T6 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	188
A.11	Plots showing altitude as a function of time, the solid curve marks the altitude of T6 above the horizon and the dashed curve marks the altitude of the moon (taken from http://catserver.ing.iac.es/staralt/).	189
A.12	A figure showing the external structure (left) of the 1.0-m telescope and it's optical diagram (right) showing how the light beam is being reflected until it reaches the detector.	197
A.13	A figure showing the Acquisition Box of the 1.0-m telescope within which the light can either be allowed to pass to the science detector or be reflected to the A/G camera. (http://www.sao.ac.za/facilities/telescope/1.0-m/).	199
A.14	An image showing the interior appearance of the warm room in the 1.0-m telescope.	200
A.15	A figure showing the interface of the TCS for controlling and monitoring the status of the telescope during the observations (Credit: Potter S. & Worters H.; TCS user's guide; version 3; 2010).	202

A.16 A figure showing how the finding chart (left) of the observed target compares to the image readout (right) from the science CCD camera. It can be seen from the images that left-right of one image is right-left of the other and top-bottom of one image is bottom-top of the other image. . 205



List of Tables

1.1	A table showing the magnitudes of ν Eridani, a β Cephei star which was discovered by Walker (1952).	9
1.2	A table showing the properties of the SITe CCD camera used during the OGLE II observation campaign.	47
1.3	A table showing the results from the OGLE II observation campaign. .	48
1.4	A table showing the properties of the “second generation” CCD mosaic camera used on the OGLE III observation campaign.	48
2.1	A list of parameters used to extract the time series data from the OGLE II database for the selection of β Cephei stars.	50
2.2	A table showing properties of six candidates selected for the November 2010, February 2011 and March 2011 observation campaigns. These properties are positions in right ascension and declination, and their photometric indices. .	71
2.3	A table showing the frequency, amplitude and FAP of the six β Cephei candidates selected for observations.	73
2.4	A table showing the properties of the STE4 CCD, it can be seen that some of the properties are specific to the 1.0-m telescope (those written 1.0-m in brackets) (see http://www.sao.ac.za/facilities/instruments/sao-ccd). . . .	75
2.5	A table showing integration time used for each Johnson’s filter employed during November 2010 observation run.	76

2.6	A table showing the starting weather conditions during observations of November 2010 targets. The targets were given the names (T1, T13) during the target selection, however their OGLE names are: T1 = LMC-SC3-i-108137 and T13 = LMC-SC13-i-218743 (see http://wasp.astro.keele.ac.uk/live/) . . .	77
2.7	A table showing the filter sequence used during afternoon sessions for taking flat field frames, columns 2 and 3 show integration time per filter and binning, respectively.	78
2.8	A table showing the coordinate differences between target stars and close-by stars whose finding charts were used to find the correct fields during observations of T13 and T1.	84
2.9	A table showing the weather conditions experienced during the February 2011 observation run, and the third column shows the number of hours spent on observing. It should be noted that by high humidity it means the humidity > 90 % and $T_{ext} - T_{Dew} < 1.5$ °C (see http://wasp.astro.keele.ac.uk/live/). . .	85
2.10	A table showing weather conditions for the third observation run during observation of T6 (LMC-SC13-i-142737) whose coordinates are $\alpha(2000.0) = 05$ 06 27.1 and $\delta(2000.0) = -69$ 06 17.4. Column 4 shows the weather during the beginning of the night (see http://wasp.astro.keele.ac.uk/live/).	87
2.11	A table showing the photometric properties of T1, T6 and T13. These were taken from the OGLE II database (see http://ogledb.astrouw.edu.pl/). . .	89
3.1	A table showing DoPhot classification types and the objects for each type.	100
4.1	A table showing the identities of stars from DoPhot reduction outputs and their corresponding identities from ISIS reduction outputs.	110
4.2	A table showing the coordinates and distance ratios of the stars identified in the LMC-T1 field and its corresponding OGLE II map. The last two columns show the distance ratios presented in equation (4.3).	129
4.3	A table showing the coordinates and distance ratios of the stars identified in the LMC-T13 field and its corresponding OGLE II map. The last two columns show the distance ratios presented in equation (4.3).	129

4.4	A table showing a list of OGLE II stars used for determining the standardizing factors for the B magnitudes of stars in LMC_T1. Column 2 contain magnitudes of stars we observed at Sutherland, column 3 contain the magnitudes from OGLE II observations and column 4 are corrected magnitudes.	133
4.5	A table showing a list of OGLE II stars used for determining the standardizing factors for the V magnitudes of stars in LMC_T1. Column 2 contain magnitudes of stars we observed at Sutherland, column 3 contain the magnitudes from OGLE II observations and column 4 are corrected magnitudes.	134
4.6	A table showing a list of OGLE II stars used for determining the standardizing factor for I magnitudes of stars in LMC_T1. Column 2 contains magnitudes of stars we observed at Sutherland, column 3 contains the magnitudes from OGLE II observations and column 4 are corrected magnitudes.	135
4.7	A table showing a list of OGLE II stars used for determining the standardizing factors for B magnitudes of stars in LMC_T13. Column 2 contain magnitudes of stars we observed at Sutherland, column 3 contain the magnitudes from OGLE II observations and column 4 are corrected magnitudes.	136
4.8	A table showing a list of OGLE II stars used for determining the standardizing factors for V magnitudes of stars in LMC_T13. Column 2 contain magnitudes of stars we observed at Sutherland, column 3 contain the magnitudes from OGLE II observations and column 4 are corrected magnitudes.	137
4.9	A table showing a list of OGLE II stars used for determining the standardizing factor for I magnitudes of stars in LMC_T13. Column 2 contains magnitudes of stars we observed at Sutherland, column 3 contains the magnitudes from OGLE II observations and column 4 are corrected magnitudes.	138
4.10	A table showing standardizing magnitude factors (average difference) and their standard deviations obtained for correcting our B, V and I magnitudes in Figs. 4.17, 4.18, 4.19, 4.20, 4.21, 4.22, 4.23, 4.24.	139
4.11	A table of coefficients for the polynomial used to calculate the effective temperature as a function of colour in equation (4.6) (taken from Torres 2010).	142

4.12	A table showing candidate β Cephei stars we obtained from the LMC-T1 field. The second and third columns show the X and Y positions of the stars on our CCD images. Their right ascension and declination coordinates are shown in columns 4 and 5. Also shown on this table are V, B and I magnitudes. . .	144
4.13	A table showing candidate β Cephei stars we obtained from the LMC-T13 field. The second and third columns show the X and Y positions of the stars on our CCD images. Their right ascension and declination coordinates are shown in columns 4 and 5. Also shown on this table are V, B and I magnitudes. . .	145
A.1	A table showing the limits, in declination, of the 1.0-m telescope (http://www.sao.ac.za/facilities/telescope/1.0-m/).	198
A.2	A table showing the properties of the UCT CCD Photometer, some of the properties are specific when the Photometer is mounted to the 1.0-m tele- scope and they are shown with “1.0-m” in brackets (http://www.sao.ac.za/facilities/instruments/uct-ccd/).	201

Chapter 1

Introduction

This thesis is a contribution to the study of asteroseismology of B stars. Asteroseismology is a study of the internal structure of pulsating stars by analysing their pulsation characteristics such as frequency. From the pulsation frequencies it is possible to determine the pulsation modes of a star. The pulsation modes are important because they provide information about the internal structure of the stars. The other importance of knowing pulsation modes of pulsating stars is that they give information about how the stars pulsate. A pulsation mode in stars is defined by three quantum numbers: n , l and m . These numbers are pulsational spherical index (l) which represents the total number of nodal lines on the surface of a star, the overtone (n) which represents the number of nodal points along the radial direction as the wave propagates inside the star, and the azimuthal order (m) which is the number of nodal lines parallel to lines of longitudes on the surface of the star. Studying pulsations in stars is important because waves causing pulsations propagate all over the stellar body, and this allows us to study parts of the star that are otherwise inaccessible and hence enhances our understanding about the interiors of stars.

Our focus in this thesis is on β Cephei stars, which are members of pulsating B stars. Several types of stars show pulsations in their light curves and spectra, these different classes are shown in Fig. 1.1. This figure also shows the instability strip of β Cephei stars and other pulsating stars.

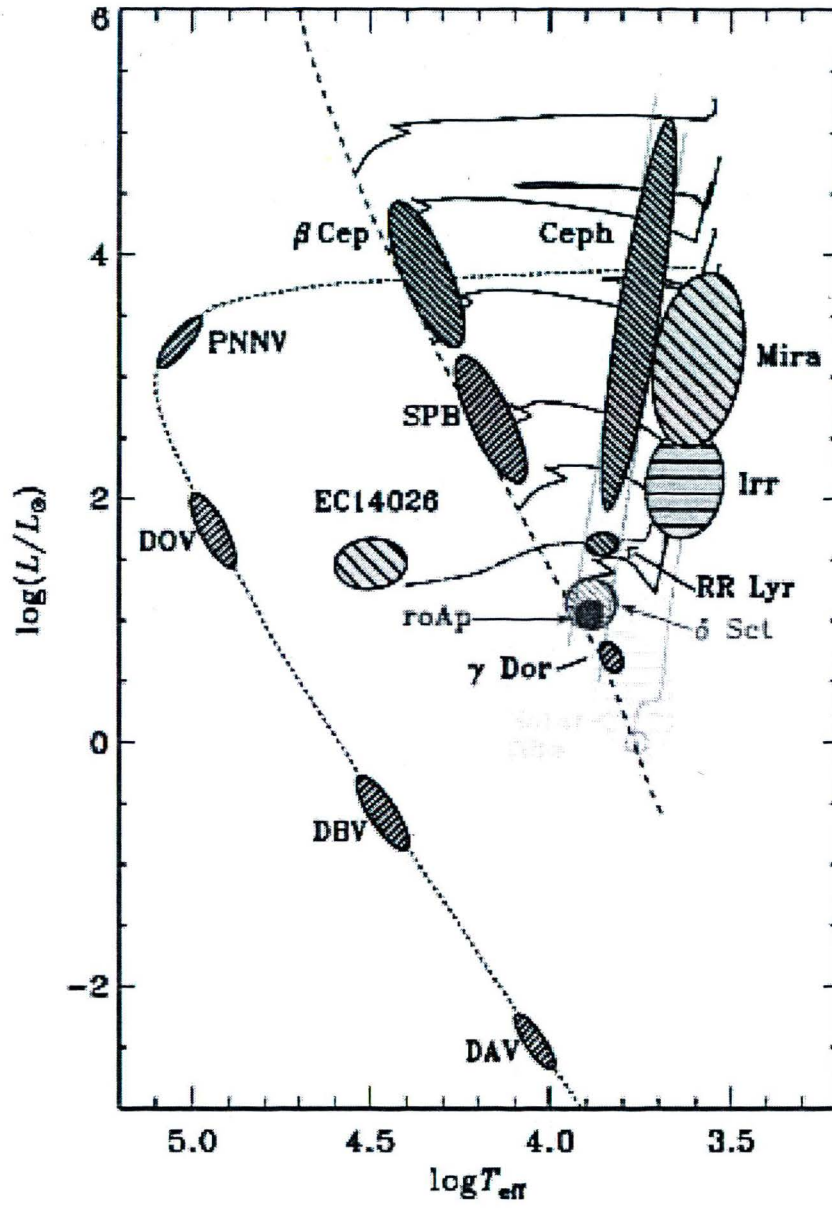


Figure 1.1: A figure showing instability strips of various pulsating stars in the HR diagram. The instability strip of β Cephei stars is on the upper main-sequence. The dashed line shows the zero-age main sequence, the continuous curves are selected evolution tracks, the dot-dashed line is the horizontal branch and the dotted curve is the white-dwarf cooling curve (taken from Christensen-Dalsgaard 1999).

The β Cephei stars are pulsating B stars, their pulsations are driven by the κ mechanism and this mechanism is strongly dependent on the abundance of iron-group ions. Theoretical models predicted that the pulsations in β Cephei stars disappear if such stars are located in a low metallicity ($Z < 0.01$) environment (Kolaczowski et al. 2006). This was confirmed by Miglio et al. (2007), where improved opacities and metal abundances were used. They found that there are no β Cephei model showing pulsations for metallicities less than 0.005. However, the OGLE data obtained for clusters in the Magellanic Clouds showed the presence of β Cephei stars (Pigulski & Kolaczowski 2002), this brought a discrepancy between theory and observations because the metallicity of Magellanic Clouds is low compared to the theoretical metallicity cut-off for β Cephei pulsations. For the Small Magellanic Cloud (SMC) the values of $0.001 \leq Z \leq 0.004$ were measured (Maeder et al. 1999), and $Z = 0.007$ for the Large Magellanic Cloud (LMC) (Martayan et al. 2008).

Since the targets observed in OGLE were observed through BVI filters only, this means that the OGLE database does not have all colours data. Our aims were to search and perform period analysis of β Cephei stars in the LMC from their multi-colour data. In order to do that, we selected B stars from the OGLE II database. A selection was made from LMC's targets which are not brighter than 14.0 magnitude and not dimmer than 16.5 magnitude. From the extracted stars, we searched for possible β Cephei stars. Having the candidate β Cephei stars, we took observations of those stars, and used the acquired data to construct the HR-diagrams. Due to time constraints we could not do frequency analysis of the reduced data to conclusively prove that we have detected new β Cephei stars.

This thesis is structured as follows, it consists of two parts. The first main part is divided into five chapters. The first chapter gives the literature review of B stars, β Cephei stars (Galactic, LMC's and SMC's), SPB stars and Be stars. The problem definition and the objectives of our study as well as a brief description of OGLE project are also given in chapter 1. The second chapter discusses the selection criteria for our

targets as well as the details of the observations we took. The reductions of the data collected as well as the computer programs used are explained in chapter 3. Chapter 4 presents the results and chapter 5 gives final remarks about the project research. The second part of this thesis is an appendix and contains the additional information such as computer codes I wrote to perform various manipulations and analysis of the data.

1.1 B stars in general

This section gives a brief classification of stars according to their temperature, it narrows to explaining B stars and their general properties.

At the beginning of the 20th century Annie Jump Cannon designed the spectral classification method for stars according to their temperature and the strength of the spectral lines (absorption and emission lines due to atoms on their atmospheres). This classification method was designed using letters O, B, A, F, G, K and M such that stars classified as O type have the highest temperature (hottest) and M stars be the coolest stars. Within this classification are B stars which are the second hottest stars after O stars. Each of these spectral types were subdivided into ten subclasses, by introducing a numerical suffix to the classification letters mentioned above. This means that, for example, B stars were subdivided into B0, B1, B2, B3, B4, B5, B6, B7, B8 and B9 with B0 stars being the hottest of all B stars. The subject of this thesis, β Cephei stars, are members of the B star class.

The B stars are hot blue and white stars whose spectra show very strong absorption lines of neutral helium (HeI). Their brightness ranges from the less luminous and less massive subdwarf and white dwarf stars, through the main sequence, core hydrogen burning stars to massive and luminous supergiant stars. The HeI absorption line is mostly strong in B2 stars, which are the third hottest B-type stars. Their spectra also show the absorption lines of neutral hydrogen (HI) and some show an emission line in H α . The B stars have a typical surface temperature range of between 10000 K for B9

stars and 30000 K for B0 stars, and they radiate most of their energy in the ultraviolet (Kaler 1997). Studies showed that luminous, main sequence B stars are not only hot but also massive. This makes them to evolve faster as well because they burn hydrogen in their cores faster.

There are various subclasses of B stars, these include normal B stars which are not pulsating. Their non-pulsating behavior is explained by Cox et al. (1992) as due to existing deficiency in iron abundance in their interiors, resulting in the reduced opacity. There are also B stars whose brightness changes with time. Among these variable B stars are Be stars, SPB stars, β Cephei stars and Bp stars. A brief summary of the properties of these B variables is given in sections 1.2, 1.5 and 1.6.

1.2 β Cephei stars

The β Cephei stars are population I B-type stars. The population I stars are the third generation of stars that have high metallicity compared to the other two older populations (i.e. II and III). The β Cephei stars used to be called β Canis Majoris stars because β Canis Major is the first star of this class to be discovered, however the usage of the name β Cephei brought a confusion with classical Cepheids and hence the name β Canis Majoris was promoted to avoid such confusion. Stankov & Handler (2005) define β Cephei stars as “*massive non-supergiant variable stars with spectral type O or B whose light, radial velocity and/or line profile variations are caused by low-order pressure and gravity mode pulsations*“. This definition was inspired by the fact that β Cephei stars are not the only known B pulsators, it was also inspired by the overlap in characteristics which occur among B-type pulsators, for instance, some Be stars and Bp stars also show the characteristics of β Cephei stars. The above definition of β Cephei stars suggests that some late O stars also show β Cephei characteristics. In fact, the observational boundary of the instability strip of β Cephei stars is still not accurately known. The β Cephei stars cannot be classified according to their pulsation modes

since no extensive mode identification has been done on these B pulsators (Stankov & Handler 2005).

The β Cephei stars have been studied since the beginning of the 20th century, and they were first defined by Lesh & Aizenman (1978) as a group of variables, having spectral type B, with light and radial velocity variations. The period of light variation ranges from three to seven hours. Lesh & Aizenman (1978) restricted the spectral type of β Cephei stars to B-type only whereas Stankov & Handler (2005) included late O stars. The other thing is the pulsation period range, which is narrower from Lesh & Aizenman (1978). These could be because the number of β Cephei stars known at the time of Lesh & Aizenman (1978) was small. However, Pigulski & Pojmański (2008) discovered 103 new β Cephei stars. This discovery doubled the number of known Galactic β Cephei stars and most of the discovered β Cephei stars showed low frequency periodic variations, which make them candidate hybrid β Cephei/SPB pulsators. Some of the β Cephei stars discovered by Pigulski and Pojmański (2008) have spectral types O8 and O9. This validated that some late O stars have β Cephei characteristics and that the spectral type range of β Cephei stars reported by Lesh & Aizenman (1978) was influenced by small number of these stars known then.

1.3 Photometric properties of β Cephei stars

This section gives the photometric properties of β Cephei stars, and because members of these stars were recently discovered outside the Milky Way (Pigulski & Kolaczowski 2002), we distinguish Galactic β Cephei stars from those in the LMC and SMC. Even though β Cephei stars were initially discovered through spectroscopic techniques, they were later studied in detail using photometry. The following subsection outlines photometric properties of β Cephei stars in the Milky Way.

1.3.1 β Cephei stars in the Milky Way

Looking at photometric searches and studies of β Cephei stars in the Milky Way, the first photometric study was the discovery of the variability on the light of β Canis Majoris by Guthnick in 1913. The light variability of β Canis Majoris was found to have the same period as the radial velocity variations and had the amplitude of 0.05 magnitude (Sterken & Jerzykiewicz 1993) which is half of the typical amplitude for variabilities in β Cephei stars. The typical value of the amplitude of light variability of β Cephei stars was reported by Lesh & Aizenman (1978) to be 0.1 magnitude, with an exception of one β Cephei star known to have the highest light variability amplitude. This is BW Vulpeculae, which has light variability amplitude of 0.2 magnitude in the V-band. This value is colour-dependent because it increases to 1.2 magnitude in ultra-violet. BW Vulpeculae has a pulsation period of about five hours and this was reported to be increasing at the rate of two seconds per century (Sterken & Jerzykiewicz 1993). However, a typical light variability amplitude of β Cephei stars reported by Lesh & Aizenman (1978) turned out to be due to an observational bias from the limited photometric accuracy at the time, this is because the lower amplitude β Cephei stars had been discovered (Pigulski & Kolaczowski 2002).

Intensive searches of β Cephei stars were conducted by Struve in the nineteen fifties, such searches were spectroscopic and their findings are discussed under spectroscopic properties of β Cephei stars (see section 1.4). From the many photometric searches of Galactic β Cephei stars, below we summarize the findings of those searches.

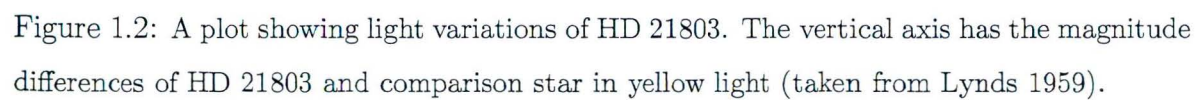
The search by Walker (1952) observed 55 early B stars in an attempt to discover those showing β Cephei characteristics. From her photometric analysis, only three stars showed significant variabilities in their light curves and one of the three was found to show β Cephei variability. This star was ν Eridani, which is a B2 star with the luminosity class of III. The brightness properties of this star, as reported by Walker (1952), are shown in Table 1.1.

Filter	Magnitude
V	3.92
U	2.82
B	3.739

Table 1.1: A table showing the magnitudes of ν Eridani, a β Cephei star which was discovered by Walker (1952).

It can be seen from Table 1.1 that ν Eridani is a bright β Cephei star. It is multi-periodic with two pulsation periods of 0.17351 day and 0.1779 day. The pulsation period was found to be changing at the rate of -1.11 seconds per century by Lesh & Aizenman (1978). The recent study of ν Eridani by Jerzykiewicz et al. (2005) showed that this star has more than two pulsation periods. The multisite photometric observing campaign, which was devoted to ν Eridani, increased the pulsation frequencies of this star to 12 high and two low frequency modes.

Seven years after Walker's (1952) discoveries, Lynds conducted another search of β Cephei stars using an EMI 6094 1 photo-electric tube, which was mounted on a 0.5-m reflecting telescope of Mount Palomar Observatory. Lynds (1959) observed only one early B-type star HD21803 and found a light variability with the period of about 4.8 hours, and an approximate brightness variation amplitude of 0.1 magnitudes. The brightness variation from night to night led Lynds (1959) to speculate that it was likely that HD21803 was a multi-periodic star. From the report of the discovery of this star, Lynds (1959) claimed that "*The light-range is seen to change from 0.110 mag, on September 15 to 0.070 mag, on September 16.*". HD21803 is a B2 star with luminosity class of IV and an absolute magnitude of -2.9 magnitudes. Fig. 1.2 shows the light curves of HD21803 as presented by Lynds (1959).



The search following that of Lynds (1959) was conducted in 1967 by Graham Hill at Kitt Peak National Observatory. Hill (1967) observed 153 early B stars using a 16-inch telescope. The observations were motivated by two reasons, firstly it was to determine the precise extent of β Cephei phenomenon and secondly, it was to discover evolutionary state of the β Cephei stars. The sample of stars observed during this search was selected from nearby stars as well as from Galactic clusters. The selected stars had spectral types ranging from O9 to B5 and the luminosity classes ranging from class II to class IV. The spectral type range was chosen to increase the chances of discovering new β Cephei stars as the then known members were having spectral types ranging from B0 to B3 (Hill 1967). Identification of β Cephei stars from observed targets was done by looking at the periodicities of less than one day in their light curves. Out of 153 observed stars, 24 were reported to be new β Cephei stars.

Another search for β Cephei stars was conducted by Mikola Jerzykiewicz in 1972. He observed the stars which were selected on the basis of their MK spectral classification, and he used a 24-inch Air Force telescope of Mauna Kea Observatory. This telescope was equipped with an EMI 62565 1 photo-electric tube. The data was taken through a Stromgren b-filter and the obtained light curves showed that the star, HR 6684 was showing a light variability with the period of 0.13989 ± 0.00001 day. Because of this period, Jerzykiewicz (1972) suggested that HR 6684 should be classified as a member of the β Cephei variables. This was because HR 6684 has the same photometric properties as δ Ceti, which is a bona fide β Cephei star. The other reason was its location on the HR-diagram, its MK spectral type of B2 IV-V placed it close to the instability strip of β Cephei stars. However, whether HR 6684 was to be considered as β Cephei star or not was a matter of spectroscopic confirmation. The light curve of HR 6684 is shown in Fig. 1.3.

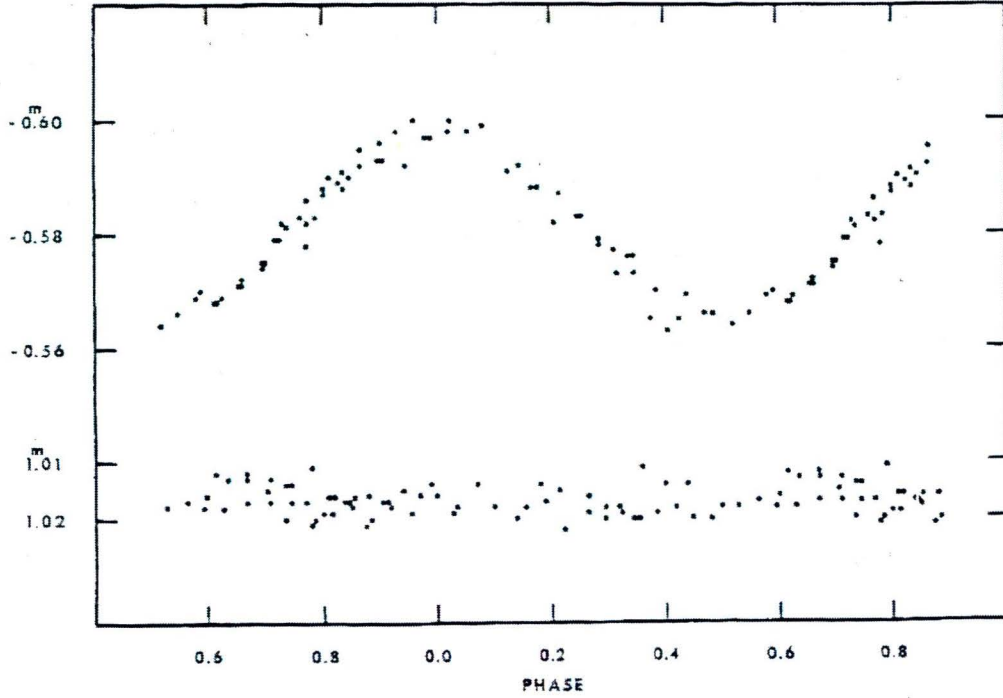


Figure 1.3: A Stromgren b-filter light curve of HR 6684 measured by Jerzykiewicz (1972), during a search for new β Cephei stars. The bottom dotted-curve is the magnitude difference of comparison stars (Taken from Jerzykiewicz 1972).

The southern hemisphere search for β Cephei stars was conducted by Balona (1977). These photometric observations were done with a 0.5-m reflecting telescope of the South African Astronomical Observatory. In this search Balona (1977) observed thirty-one candidates, eleven stars showed β Cephei characteristics. The properties of the discovered β Cephei stars are given in Balona (1977), showing spectral types ranging from O9.5 to B3, with luminosity classes ranging from III to IV. Balona (1977) plotted the HR-diagram (M_{bol} vs $\log T_{\text{eff}}$) of discovered β Cephei stars, this is shown in Fig. 1.4.

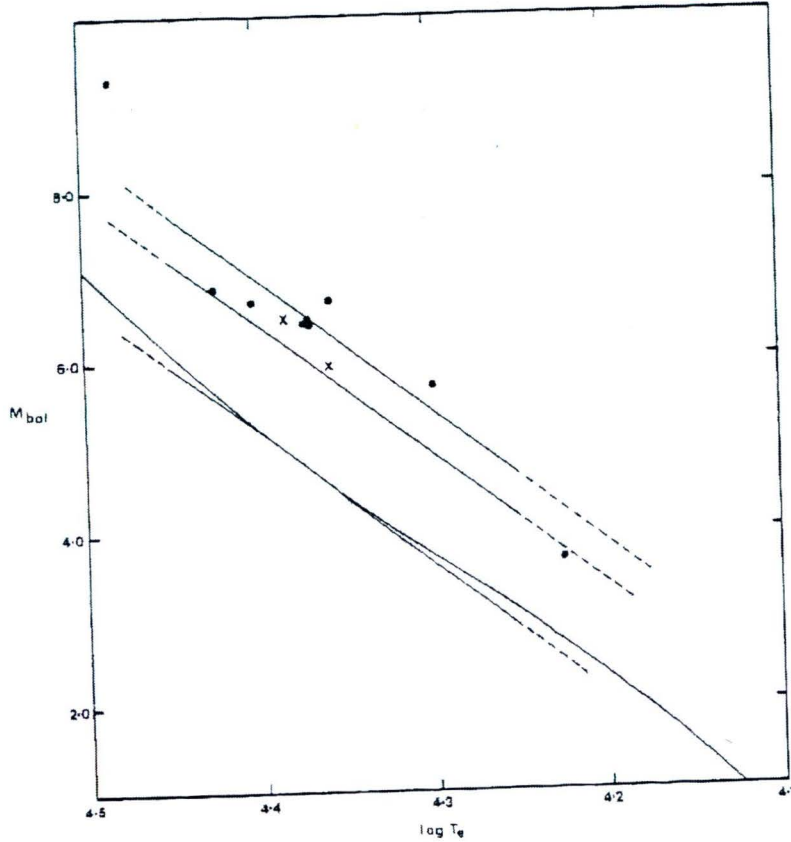


Figure 1.4: A plot showing HR-diagram of stars Balona (1977) observed. The curved line is a semi-empirical Zero-Age Main Sequence (ZAMS), dashed line next to ZAMS is a theoretical ZAMS. The upper two lines show theoretical region of post main sequence evolution (Taken from Balona 1977).

Not too long after that, Balona (1983) conducted another search in the Galactic cluster, NGC 6231. In that search, three new β Cephei stars were discovered from the list of thirty-five observed candidates. The discovered β Cephei stars have pulsation periods of 0.10 day, 0.07 day and 0.11 day (Balona 1983). These β Cephei stars were concluded not to have evolved very far from the zero-age main sequence because their host cluster is young. This conclusion posed the idea that an instability strip of β Cephei stars, which was theoretically found to be one magnitude above and parallel to zero-age main sequence, appeared to be the result of observational selection effect (Balona 1983).

Then there was another β Cephei search in Galactic cluster, NGC 4755 by Chris Koen in 1993. Three stars were found showing β Cephei characteristics. Koen (1993) found that one of the three β Cephei stars he discovered was already known and was discovered by Jakate (1979). Their three β Cephei stars F, G and I have spectral types B2 III, B0.5 V and B1 V respectively. A known β Cephei star F was found to be multi-periodic with pulsation frequencies of 4.89 and 6.16 cycles per day. The amplitude of these frequencies were reported to be 7 mmag and 6 mmag, respectively. The second β Cephei star G was also found to be multi-periodic with three frequencies given as 6.62, 6.33 and 0.21 cycles per day and the other β Cephei had just a single pulsation frequency of 5.59 cycles per day (Koen 1993).

There were also studies of β Cephei stars in binary systems, they include two eclipsing systems in Stock 14, a young open cluster. One eclipsing system consists of candidate β Cephei star HD 101794 and the other system of another candidate β Cephei star HD 101838. The light variability of these two suspected β Cephei stars was verified to be due to eclipsing binary and pulsations by Drobek et al. (2010). HD 101794 was found to have the orbital period of $P_{orb} = 1.4632$ days while HD 101838 was found to have $P_{orb} = 5.41167$ days. Drobek et al. (2010) analyzed the light curves of these stars by performing Fourier transform and used the orbital periods mentioned above to show that the pulsation frequencies of HD 101794 is $f_1 = 4.4482$ cycles per day and $f_2 = 1.8760$ cycles per day, and that for HD 101838 is $f = 3.1299$ cycles per day. Figs. 1.5 and 1.6 show the frequency spectra of these two stars.

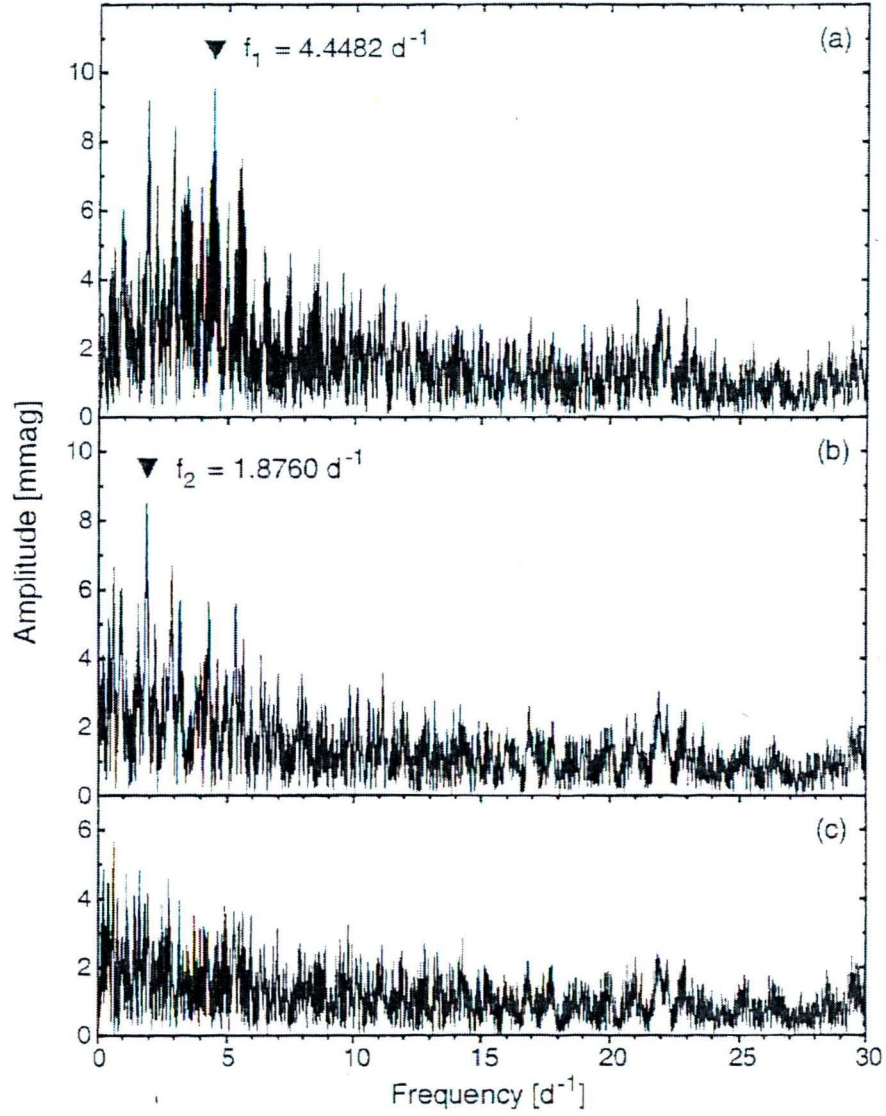


Figure 1.5: A plot showing frequency spectra of HD101794. The panels from the top show the spectra in the sequence of prewhitening, the top panel marked (a) shows the first frequency, middle panel (b) shows the second frequency with the first frequency peak removed and the last bottom panel (c) shows a spectrum with all significant peaks removed (taken from Drobek et al. 2010).

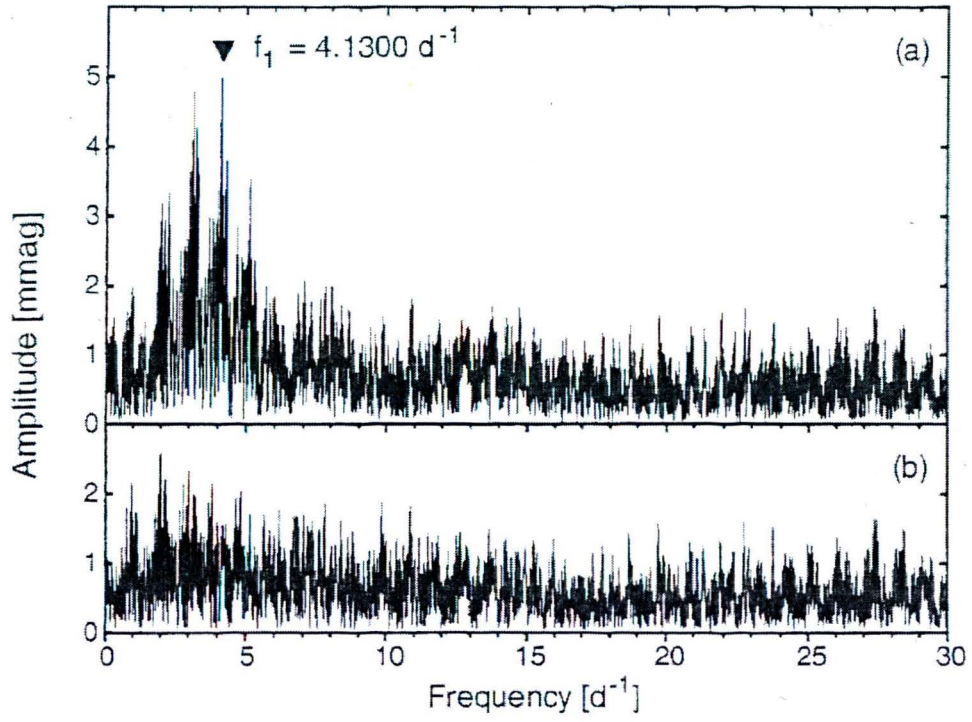


Figure 1.6: A plot showing frequency spectra of HD101838 in the sequence of prewhitening from the top spectrum, top panel (a) shows the only frequency peak found and the bottom panel (b) is the spectrum with the peak removed and mostly showing noise (taken from Drobek et al. 2010).

The recent discovery of new β Cephei stars was done by Handler & Meingast (2011). Their search was done on the galactic cluster, NGC 637. Five stars, one of which was a known β Cephei star, were found to be showing β Cephei characteristics. These stars are reported by Handler & Meingast (2011) to have made NGC 637 one of the six young clusters with significant number of β Cephei stars, the other clusters with significant number of β Cephei stars are NGC 884 (Pigulski 2004), NGC 3293 (Stankov & Handler 2005), NGC 4755 (Koen 1993), NGC 6231 (Balona 1983) and NGC 6910 (Saesen et al. 2010). The three newly discovered β Cephei stars (i.e. NGC 637 1, NGC 637 7 and NGC 637 138, which are clearly stars 1, 7 and 138 of NGC 637) by Handler & Meingast (2011) were found to be multi-periodic. Their light curves together with their prewhitened frequency spectra are shown in Figs. 1.7, 1.8, 1.9, 1.10 and 1.11. The other features shown on these figures are light curves of the two previously known β Cephei stars.

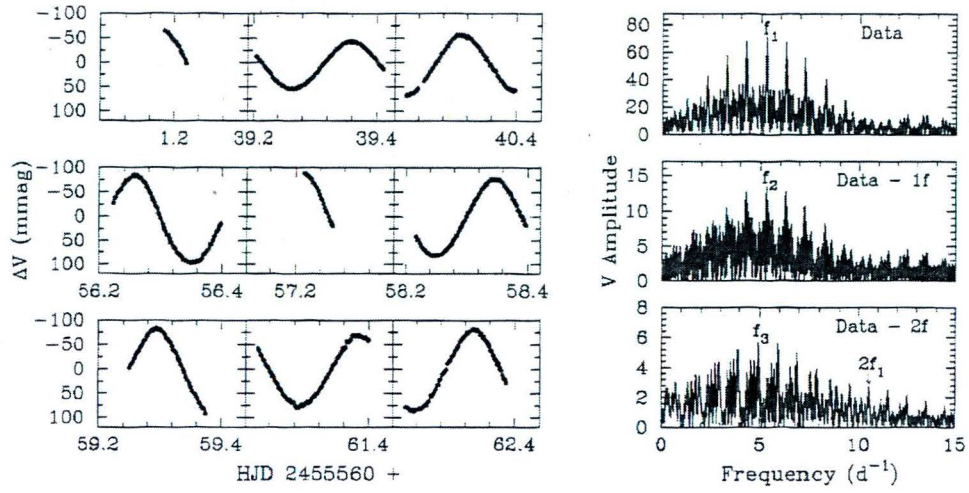


Figure 1.7: A plot showing light curve (left) and frequency spectra (right) of β Cephei star NGC 637 4. The frequency spectra are in the order of prewhitening from top to bottom (taken from Handler & Meingast 2011).

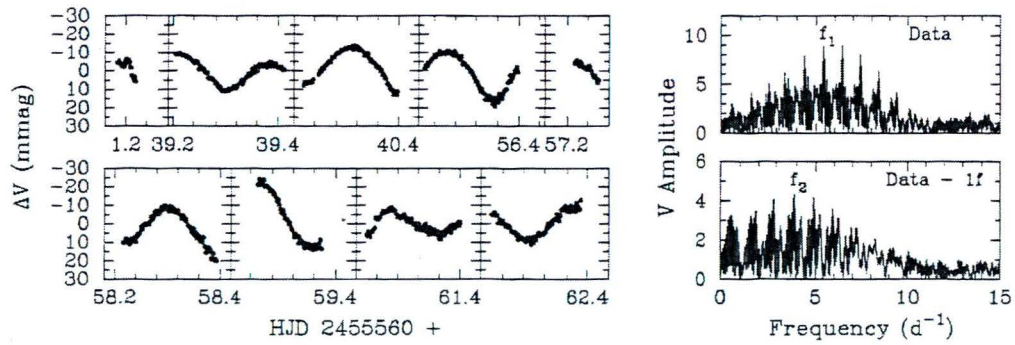


Figure 1.8: A plot showing light curve (left) and frequency spectra (right) of β Cephei star NGC 637 1. The frequency spectra are in the order of prewhitening from top to bottom (taken from Handler & Meingast 2011).

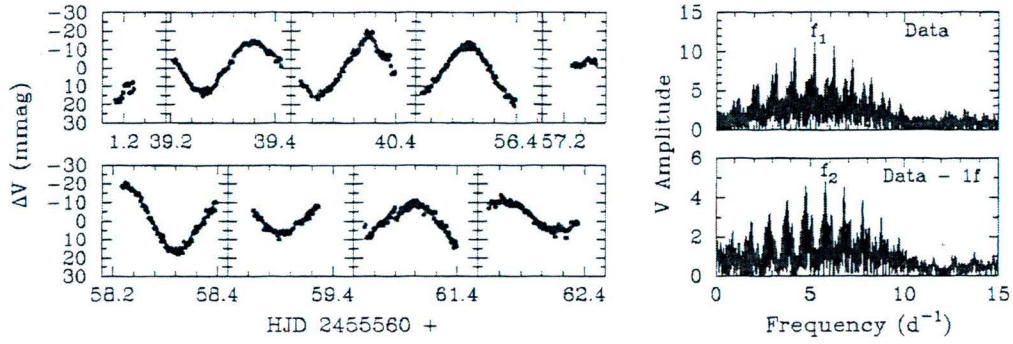


Figure 1.9: A plot showing light curve (left) and frequency spectra (right) of β Cephei star NGC 637 7. The frequency spectra are in the order of prewhitening from top to bottom (taken from Handler & Meingast 2011).

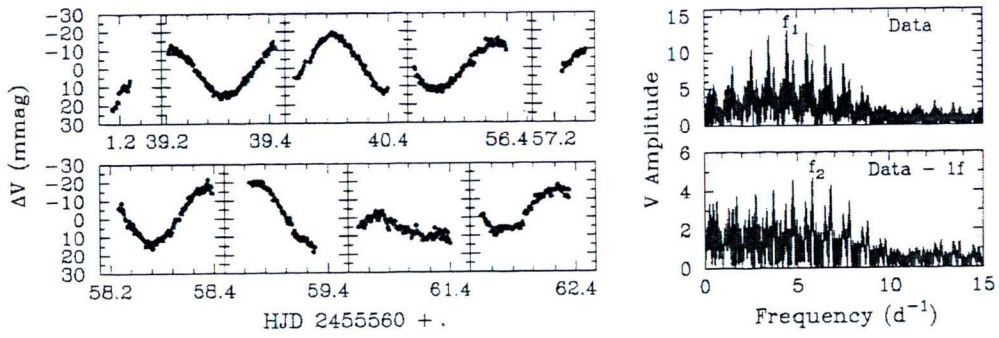


Figure 1.10: A plot showing light curve (left) and frequency spectra (right) of β Cephei star NGC 637 138. The frequency spectra are in the order of prewhitening from top to bottom (taken from Handler & Meingast 2011).

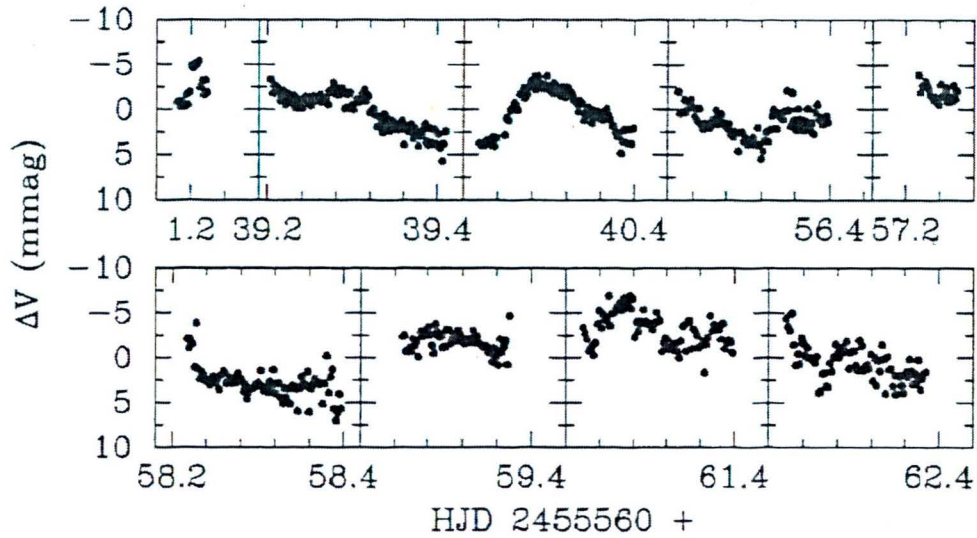


Figure 1.11: A plot showing light curve of suspected β Cephei star NGC 637 6 (taken from Handler & Meingast 2011).

The β Cephei stars are among the most studied B stars, and there is plenty of literature on these stars. To conclude the photometric studies of Galactic β Cephei stars, the following subsection gives a summary.

1.3.2 Summary of photometric properties of Galactic β Cephei stars

From the analysis of data published in the past 100 years, Stankov & Handler (2005) were able to come up with the general properties of confirmed β Cephei stars. All analyzed β Cephei stars have the luminosity classes ranging from I to V, with B1 stars dominating at the luminosity classes of III and V, followed by B2 stars at the luminosity classes of III and IV. This is presented as a 3-D histogram of spectral types and luminosity classes by Stankov & Handler (2005), and such histogram is shown in Fig. 1.12.

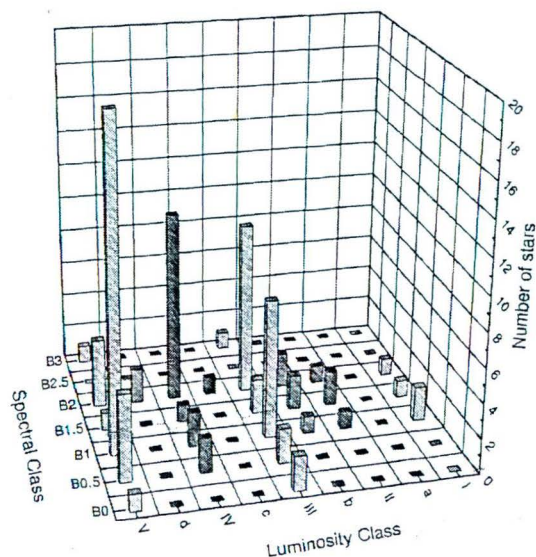


Figure 1.12: A 3D plot showing the distribution of β Cephei stars according to their spectral class and luminosity class (taken from Stankov & Handler 2005).

It can be seen from Fig. 1.12 that all β Cephei stars have spectral types ranging from B0 to B3, where B1 types with luminosity class V are the majority as it can be seen from Fig. 1.12. The luminosity class reported is in agreement with the one on the review of β Cephei stars by Sterken & Jerzykiewicz (1993), however the spectral types differ. Sterken & Jerzykiewicz (1993) reported the spectral type range from B0.5 to B2. This could be due to the small sample, consisting of 21 confirmed β Cephei stars, from which Sterken & Jerzykiewicz (1992) made their conclusions as compared to the sample of 93 confirmed β Cephei stars from which Stankov & Handler (2005) made their conclusion. The spectral range differences could also be due to the fact that Stankov & Handler (2005) re-analyzed some results from previous publications, and their analysis showed that 61 B stars reported as β Cephei stars were actually not.

The period analysis of 93 β Cephei stars by Stankov & Handler (2005) showed that the photometric periods of those stars range from 0.0667 day to 0.32 day. The histogram of all 93 β Cephei stars shows that most of them have periods between 0.16 day and about 0.175 day, this is shown in Fig. 1.13.

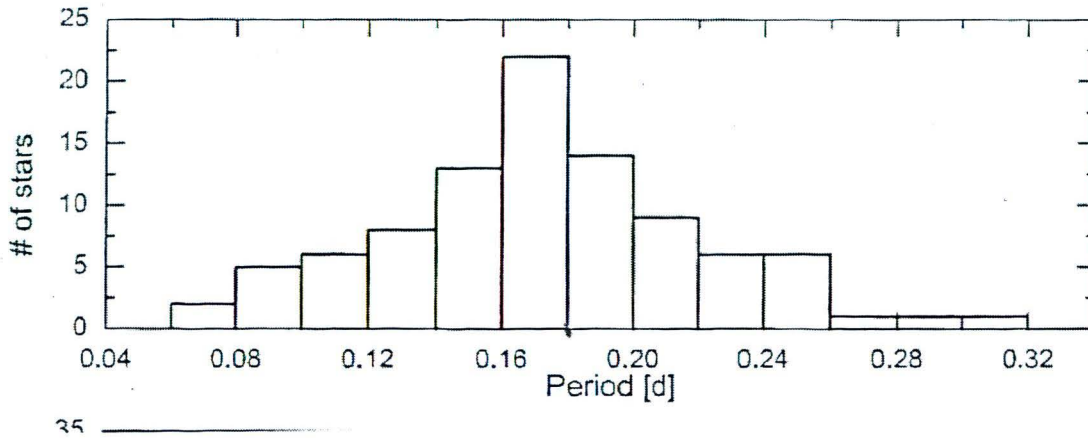


Figure 1.13: A plot showing the distribution of Galactic β Cephei stars with pulsation period by 2005 (taken from Stankov & Handler 2005).

The β Cephei stars reported on this 2005 review were brighter than 12 magnitude, and most of them have V-magnitudes between 9 and 10. However, there was one β Cephei star with 15 to 16 magnitude as it can be seen in Fig. 1.14.

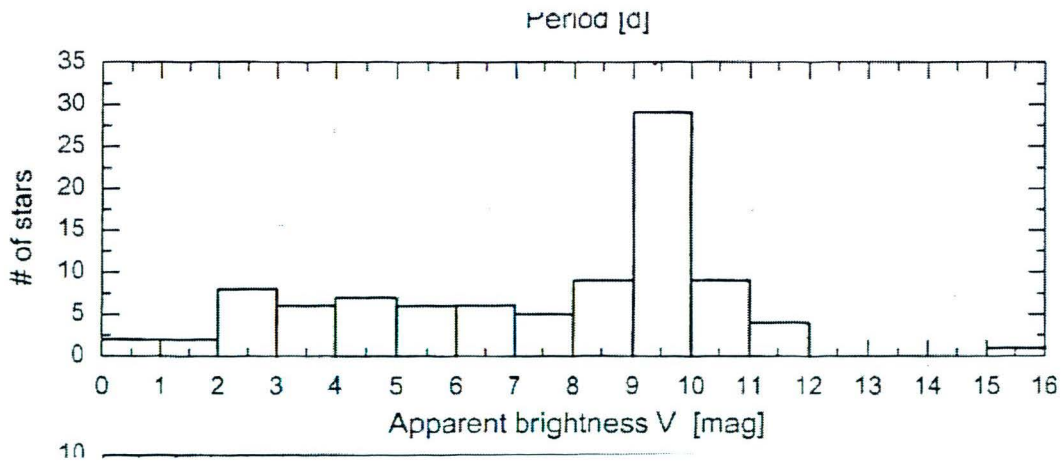


Figure 1.14: A plot showing the distribution of Galactic β Cephei stars in V magnitudes by 2005 (taken from Stankov & Handler 2005).

By the end of 2005 there were two known β Cephei stars with the modes of high-degree l . The β Cephei stars with these modes were reported by Stankov & Handler

(2005) to have light variability which is difficult to detect. These stars were classified as β Cephei stars from their line profile variability and the basic behaviour similar to those of β Cephei stars.

In conclusion, it can be seen that these stars are early B-type stars with MK classification ranging from B0 to B3 and luminosity class ranging from III to V, with few β Cephei stars having luminosity classes of I and II (see Fig. 1.12). They pulsate with high overtone pressure modes (p-modes) and gravity modes (g-modes) with pulsation period ranging from 1.6 hours to 7.7 hours (Stankov & Handler 2005). This period range is reported differently by few authors (see for example Aerts, Christensen-Dalsgaard & Kurtz 2009). Most, if not all, known β Cephei stars have masses between $8 M_{\odot}$ and $18 M_{\odot}$. It can be seen from figure 6 of Stankov & Handler (2005) that there are few known β Cephei stars with masses greater than $18 M_{\odot}$. Most known β Cephei stars have light variability with an amplitude less than 0.1 magnitude, with an exception for BW Vulpeculae.

It was not until in 2002 when the very first extragalactic β Cephei stars were discovered. The following subsection gives the details of the discovery of these extragalactic β Cephei stars.

1.3.3 β Cephei stars in Magellanic Clouds

The very first extragalactic β Cephei stars were discovered in the LMC. Pigulski & Kozłowski (2002) did a thorough analysis of light curves of B stars from the OGLE II database, supplemented by a MACHO database, of the LMC bar. From their analysis they discovered the very first three extragalactic β Cephei stars with masses ranging between $10 M_{\odot}$ and $20 M_{\odot}$. This discovery was not expected because of the low metallicity of the LMC as compared to metallicity required for β Cephei pulsations. However, the locations of these three β Cephei stars on the LMC were reported by Pigulski & Kozłowski (2002) to be suitable for these stars to have higher than average metallicities. These β Cephei stars were found to be located very close to massive and young LMC

associations (i.e. LH 41, LH 59 and LH 81). The two of these three β Cephei stars were found to be multi-periodic with pulsation frequencies of $f_1 = 3.49709090$ cycles per day and $f_2 = 3.685343$ cycles per day for OGLE052809.21-694432.1, and $f_1 = 3.495009$ cycles per day, $f_2 = 3.816132$ cycles per day, $f_3 = 1.684015$ cycles per day, $f_4 = 2.005502$ cycles per day and $f_5 = 5.179046$ cycles per day for OGLE051841.98-691051.9. The last β Cephei star, OGLE053446.82-694209.8 has one pulsation frequency of $f = 4.05160$ cycles per day. All these three β Cephei stars are fairly faint and they have the V-magnitudes of 16.691, 16.748 and 14.327 for OGLE053446.82-694209.8, OGLE052809.21-694432.1 and OGLE051841.98-691051.9 respectively. One of the three discovered extragalactic β Cephei stars showed to have spectral type outside of the range of B0-B3. This is shown on the colour-magnitude diagram of Pigulski & Kolaczowski (2002) and such diagram is reproduced in Fig. 1.15. The three β Cephei stars are shown as filled circles and from their V-magnitudes mentioned above, it can be seen that OGLE051841.98-691051.9 is the one which is outside the spectral range of Galactic β Cephei stars.

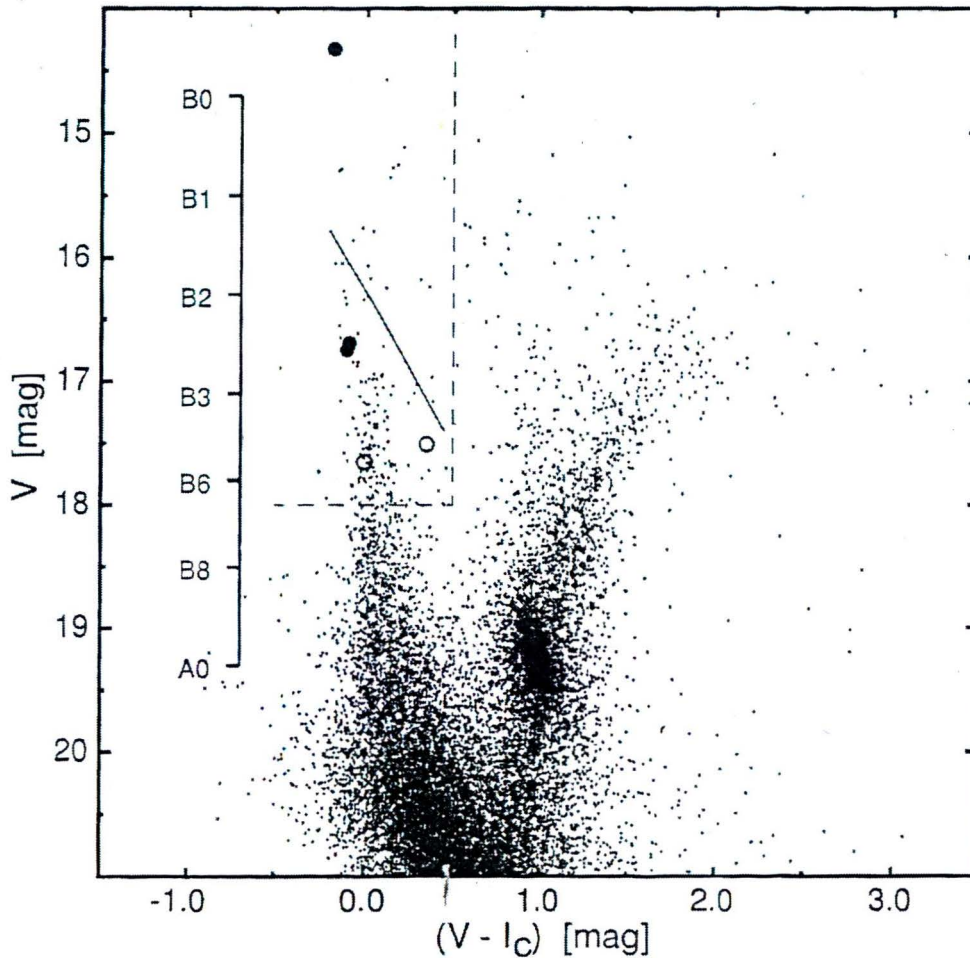


Figure 1.15: A plot showing colour-magnitude relation of the sample of stars from the LMC. The three discovered extragalactic β Cephei stars are shown as filled circles, empty circles are suspected β Cephei stars (Taken from Pigulski & Kolaczowski 2002).

The discovery of only three β Cephei stars in the sample of 5204 early B-type stars in LMC area extracted by Pigulski & Kolaczowski (2002) led them to conclude that, their finding was in agreement with theoretical predictions. Theoretically, pulsations in B stars fail to exist if such stars are located in a low metallicity environment. But it turned out that such discovery of only three β Cephei stars was due to a small sample, because not long after the first three extragalactic β Cephei stars were discovered, Pigulski & Kolaczowski (2004) re-analyzed the OGLE II time-series data and discovered 64 β Cephei stars in LMC. The same data-extraction parameters (i.e. V magnitude

less than 18 magnitude and (V-I) colour less than 0.5 magnitude) were used in both 2002 and 2004 searches, but the only difference was that in 2002 search. Pigulski & Kolaczowski (2002) analyzed a sample of 5204 stars whereas in 2004 the sample was bigger, it consisted of 75,000 stars. Out of those 64 β Cephei stars discovered, Pigulski & Kolaczowski (2004) found that 31 β Cephei stars have single mode, and the rest have more than one mode. Their colour-magnitude diagram (see Fig. 1.16) shows that all discovered 64 β Cephei stars have V-magnitudes ranging from 14 to 17.5 magnitudes. Most of these LMC β Cephei stars were found to have pulsation periods longer than the Galactic β Cephei stars, the period median of 0.27 day was calculated for LMC members as compared to period median of 0.17 day for Galactic β Cephei stars. Pigulski & Kolaczowski (2004) suggested that the longer periods of LMC members could be due to the presence of low-order gravity modes in these stars. This would imply that those stars could be regarded as β Cephei as well as SPB stars. The colour-magnitude diagram of 64 β Cephei stars discovered in 2004 is shown in Fig. 1.16.

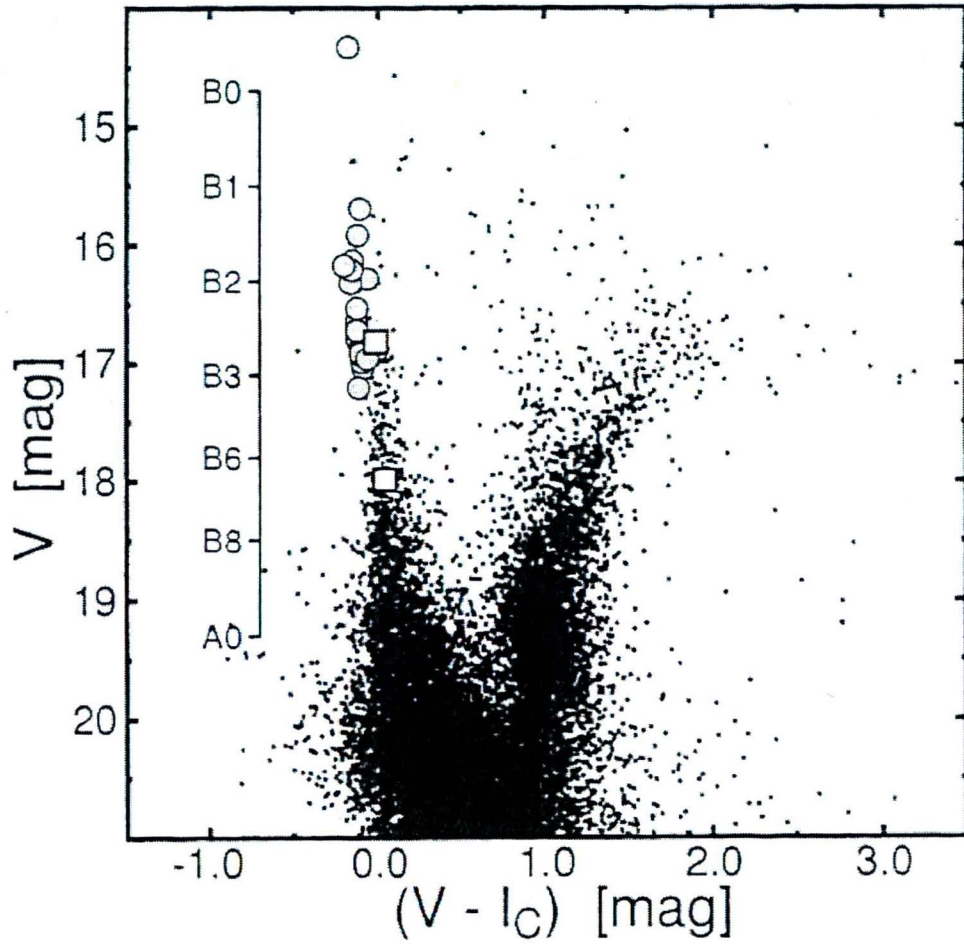


Figure 1.16: A plot showing colour-magnitude diagram of the sample of LMC stars, from which 64 β Cephei pulsators were discovered. These 64 β Cephei stars are represented by circles and suspected β Cephei stars are represented by squares (taken from Pigulski & Kolaczowski 2004).

Kolaczowski et al. (2006) discovered 92 β Cephei candidates in the LMC from the sample of 215,000 stars with V magnitudes less than 18.5 magnitude and colour (V-I) less than 0.5 magnitudes. These stars were extracted from the OGLE II database in 32 fields of both the LMC and the SMC. In SMC, six β Cephei stars were discovered. Most of these 92 LMC and 6 SMC β Cephei stars were found to be multi-periodic and colour-magnitude diagrams of these stars are shown in Fig. 1.17.

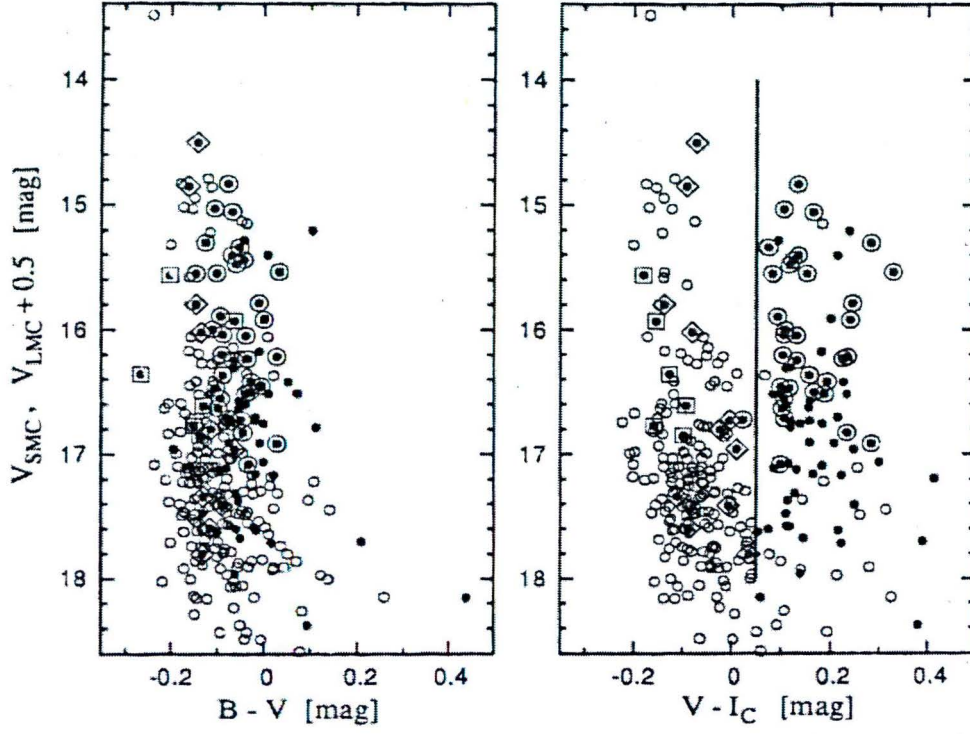


Figure 1.17: A plot showing colour-magnitude diagrams (i.e. (B-V) on the left panel and (V-I) on the right panel) of a sample of LMC and SMC stars. The discovered β Cephei stars are shown with open circles and dots in squares for LMC and SMC, respectively. These diagrams also show discovered Be stars (encircled dots) and the vertical line on the right panel separates Be stars and non-Be stars (taken from Kolaczowski et al. 2006).

Since β Cephei stars were first discovered using spectroscopic techniques, the following section gives the spectroscopic properties of Galactic members of these variables. Spectroscopic information of β Cephei stars in the Magellanic Clouds is not yet available.

1.4 Spectroscopic properties of β Cephei stars

It was mentioned above that the very early studies and observations of β Cephei stars were done using spectroscopic techniques. The discovery of the first Galactic β Cephei star was done using the Bruce Spectrograph of Yerkes Observatory at Wisconsin, United States of America. From the six spectrograms (i.e. time varying spectra), Frost (1902) discovered that radial velocity of β Canis Majoris is a variable, but the variability period could then not be determined due to inadequate measurements. However, Frost (1902) predicted that the variability period was short from two spectrogram plates measurements. Four years later Frost (1906) confirmed his short variability period prediction by using measurements from eight spectrogram plates to measure the period of approximately 4.5 hours. Using the spectrum obtained from the six spectrograms mentioned above, Frost (1902) identified the absorption lines due to Oxygen, Silicon, Helium and Magnesium, which are certainly the elements found on the surface of most, if not all, β Cephei stars. These lines were used by McNamara et al. (1955) to determine the radial velocity of BW Vulpeculae. The velocity curve of the β Cephei star reported by Frost (1906) is shown in Fig. 1.18. This plot was used to measure the above mentioned approximate variability period of β Canis Majoris.

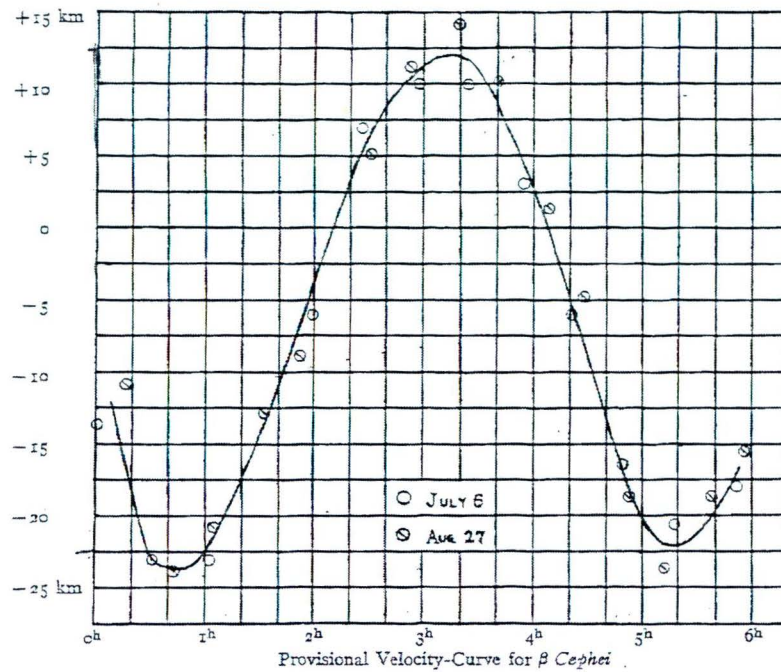


Figure 1.18: A figure showing the *provisional* velocity-curve for β Cephei discovered by Frost (1902). The circles are the data points and the curve is the fit to the data points. One should note a mistake Frost (1906) did by expressing velocity in units of a distance in vertical axis.

It can be seen from Fig. 1.18, as it was mentioned by Frost (1906), that the curve is symmetrical and its amplitude is approximately 34 km/sec, this was estimated by taking the difference of the values in the vertical axis of the trough and the peak.

Frost (1906) suggested that β Canis Majoris could be a spectroscopic binary with the orbital radius of 45,000 km. It was in 1913 when P. Guthnick discovered that this star also shows light variability with period similar to a radial velocity variability period. The discovery of light variability falsified the spectroscopic binary hypothesis because the light curve did not resemble those of ordinary eclipsing variables. Otto Struve, who did a significant input on the study of β Cephei stars, also argued that the spectroscopic binary hypothesis could not be true. Struve (1955) reported that *For spectroscopic binaries of types O, B and A, the amplitude of the velocity curve increases rapidly, on the average, with decreasing period, in accordance with Kepler's third law.*

But this increase continues only until a period of about 1.3 days is reached. For still shorter periods the amplitude again decreases. This can mean only one thing: β Cephei and all similar stars are not true binaries. Struve (1955) suggested that β Cephei behavior resembles those of Cepheids with just small difference on the range of light variations.

The spectroscopic search and study of β Cephei stars became intensive from 1920's and by mid-1955 there were ten confirmed β Cephei stars. Those stars have spectral types ranging from B1 II to B2 IV. Their radial velocities range from 6 km/sec to 150 km/sec. They also show line profile variation with periods ranging from $4^h 1^m$ to $6^h 2^m$ (Struve 1955).

Struve et al. (1953) obtained 201 spectrograms of β Canis Majoris. Using the measurements from those spectrograms, he investigated the change in the period of the line profiles and the amplitude of the radial velocity of β Canis Majoris. From the velocity-curve obtained it was found that there was a large scatter of individual points, this large scatter was reported to suggest the presence of irregular changes in the velocity-curve. The conclusion reached was that the amplitude of radial velocity has decreased as compared to the previously obtained values. As for the period of a line profile, Struve et al. (1953) found that there were certain changes between successive cycles. That allowed him to conclude that the velocity-curves of β Canis Majoris have irregular variations which turned out to cause distortion on the velocity-curve.

Earlier, under photometric properties of β Cephei stars it was mentioned that BW Vulpeculae has a higher light variation amplitude, and some of its photometric properties were also stated. McNamara et al. (1955) did a spectroscopic study of this star. Prior their study, they already knew from unpublished work of Petrie in 1937 that BW Vulpeculae has a large total amplitude of the velocity-curve, which they reported to be 150 km/sec. They also knew that this amplitude increases at the rate of 0.7 km/sec/year. The other things which were known, were an increasing spectral line

variation period at the rate of 3.7 sec/century, the small change in the shape of the velocity-curve on its descending branch, and the change in the shape of spectral line with period similar to the period of velocity-curve. McNamara et al. (1955) planned to study the behavior of the velocity-curve and also of the nature of variations in the line profiles of BW Vulpeculae. From their results they found that the total amplitude of the velocity-curve was changing from one cycle to another. Their velocity-curve also showed the beat phenomenon and they also noticed the rapid change in radial velocity. That change was of the order of 80 km/sec in 15 minutes. On the behavior study of the line profile, it was found that the sharpness of the line profile is more defined on the ascending branch of the velocity-curve as compared to the descending branch, and also that the sharpness coincides with the stillstand stage of the velocity-curve.

A standstill on the velocity-curve or light curve is a bump-like feature. Pesnell & Cox (1980) investigated resonance phenomena and non-linear pulsations mechanisms in an attempt to model and find the cause of the bump in the light curve and radial velocity curves of BW Vulpeculae. They defined the resonance mechanism by the ratio of the periods of a second overtone radial and radial fundamental modes, and it was expected that the model of BW Vulpeculae would have the period ratio less than one half for resonance to be responsible for the bump. However, they found that the period ratio was shortened much less than one half, hence the resonance was ruled out of being causing the bump. Investigations on Non-linear pulsations mechanism, which was defined by theoretical radial velocity amplitude, showed the presence of the bumps which Pesnell & Cox (1980) reported to be due to the rapid increase in the amplitude that resulted when the theoretical amplitude was increased to obtain a full amplitude oscillation.

The time lag in the radial velocities from the hydrogen lines of the β Cephei stars, β Canis Majoris and 16 Lacertae, were first discovered by A. Van Hoof in 1953. This time lag phenomenon was also found in other β Cephei stars; BW Vulpeculae, σ Scorpii and 12 (DD) Lacertae. Struve (1955) showed this phenomenon on the plot of radial velocity

against time and such plot is shown in Fig. 1.19. The velocity-curve determined from hydrogen lines maintained the same shape as the velocity-curve from other elements, except that it has been displaced to the right.

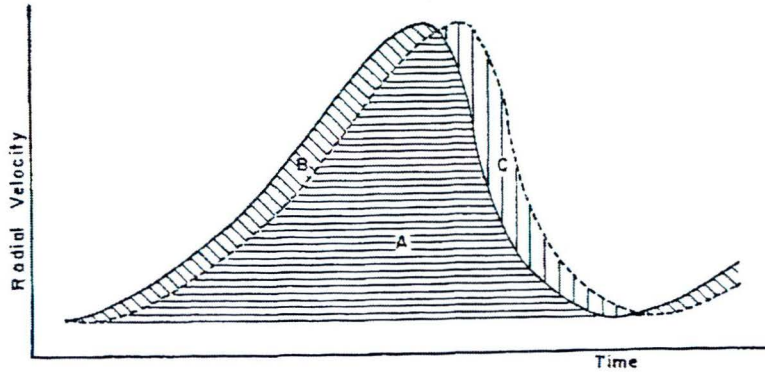


Figure 1.19: A figure showing a continuous curve which represents the radial velocities of a pulsating star from Si IIII, O II, C II. The dashed curve is the radial velocity curve for H I (taken from Struve 1955).

Struve (1955) also noticed that the time lag on the velocity-curves of BW Vulpeculae, σ Scorpii and 12 (DD) Lacertae was clearly noticeable only upon the descending branch and suggested that if there was a discontinuity in the velocity-curves then the time lag would be due to a “*delay of the components of the double hydrogen lines to reach similar ratios of intensity as do the line of Si III, O II, etc*” (Struve 1955). For the case where a discontinuity was ignored, Struve (1955) could not understand the cause of the time lag except that the velocity-curve looked like in Fig. 1.20 where only the descending branch of the velocity-curve shifted to the right.

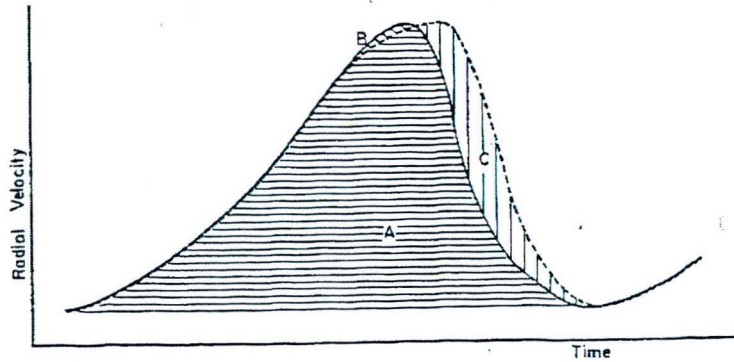


Figure 1.20: A figure showing a continuous curve representing the radial velocity of Si III, O II, C II and a dashed curve representing the radial velocity of He I (taken from Struve 1955).

Trying to understand the cause of the time lag for continuous velocity-curve case, Struve (1955) investigated the behavior of the Helium lines. This was done by determining the radial velocities due to all the lines in the spectra of σ Scorpii and 12 (DD) Lacertae, and instead of studying the behavior of the Helium lines throughout the cycle, Struve (1955) calculated the velocity differences (i.e. He I velocity minus O II, Si III and Mg II velocities; and He I velocity minus O II, Si III and Mg II velocities). Struve (1955) found that the average velocity differences for both He I and He I were positive. The He I lines were further subdivided according to their strength and the outcome was that the average velocity difference of strong lines was larger than of weak lines. This led Struve (1955) to be convinced that the amount of the departure depends on the total equivalent width of the line.

The line profile variation (LPV) of 26 bright β Cephei stars was studied by Aerts & De Cat (2003). On their study they performed the frequency analysis and mode identification in order to come up with the temporal behavior of the pulsations, and also to derive the geometrical configuration of the modes and velocity parameters. From frequency analysis, Aerts & De Cat (2003) found that their sample of β Cephei stars showed a slow to moderate rotational behavior, except two, i.e. HD52918 with $v \sin i = 274$ km/sec and HD116658 with $v \sin i = 160$ km/sec (see Aerts & De Cat 2003).

From mode identification analysis, Aerts & De Cat (2003) found that there was no clear relation between the degree of mode and the rotational velocity. The detected modes on various β Cephei stars used on this LPV study were found to have a wide variety in frequencies, radial velocity amplitude and wave numbers. However, the authors realized that better mode identification methods and more extended data sets were required in order to improve the quality of their findings.

Daszyńska-Daszkiewicz & Niemczura (2003) did the study of ultra-violet spectra of β Cephei stars observed by International Ultraviolet Explorer (IUE) satellite. Their main aim was to determine the stellar parameters (i.e. effective temperature (T_{eff}), metallicity [$\frac{m}{H}$], stellar diameter (θ) and interstellar extinction $E(B-V)$). The usage of UV spectra was motivated by the fact that, for early type stars, including β Cephei stars, most of the flux is emitted in UV spectral region. Furthermore, the dominating elements in the spectrum are the iron-group elements, these elements are of importance in the mechanism driving pulsations of β Cephei stars, and this mechanism is explained in section 1.7. Daszyńska-Daszkiewicz & Niemczura (2003) used an algorithmic procedure of fitting theoretical flux distribution to low-resolution spectra of IUE β Cephei stars and also on optical spectrophotometric observations to derive the above mentioned parameters. On their analysis they used the stellar evolutionary model equation:

$$\log g = -12.5894 + 4.4810 \log T_{\text{eff}} - 0.7870 \log \frac{L}{L_{\odot}} \quad (1.1)$$

where $\log g$ is the logarithm of gravity and $\log(\frac{L}{L_{\odot}})$ is the logarithm of the luminosity of the star in units of solar luminosity. To calculate $\log(\frac{L}{L_{\odot}})$, Hipparcos parallaxes were used and a plot of $\log g$ against $\log(T_{\text{eff}})$ was made. This plot is shown in Fig. 1.21 for different metallicities (left panel) and different clusters (right panel).

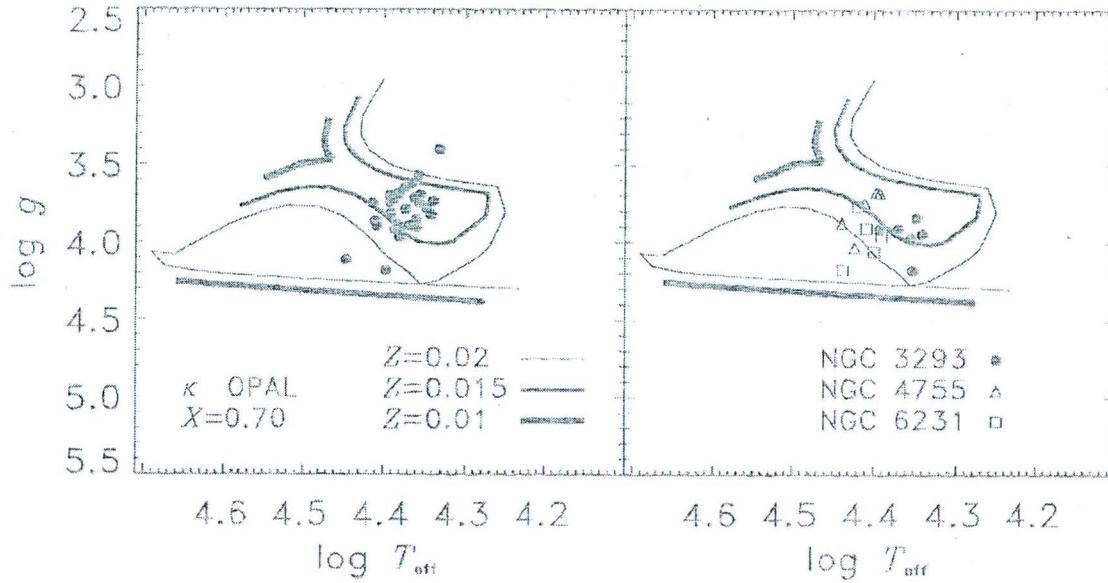


Figure 1.21: A figure showing theoretical instability strips and observational data of β Cephei stars in the $\log g$ vs $\log(T_{\text{eff}})$ diagram. The left panel is for different metallicities while the right panel is for different clusters (taken from Daszyńska-Daszkiewicz & Niemczura 2003).

From Fig. 1.21, Daszyńska-Daszkiewicz & Niemczura (2003) determined the metallicity mean value for all stars on their sample to be -0.13 ± 0.03 dex, where each of the clusters had a mean value of 0.05 ± 0.05 dex for NGC 3293, -0.43 ± 0.05 dex for NGC 4755 and -0.01 ± 0.06 dex for NGC 6231. For the field stars the mean metallicity value was found to be -0.14 ± 0.03 dex. In an attempt to check whether there was any correlation between obtained metallicities, interstellar extinction and effective temperature, they found that there was a small correlation between these parameters. This was reported to be implying that $[\frac{m}{H}]$ can be reliably derived from the best-fit procedure.

Recently Stateva et al. (2010) presented their preliminary results of the observations of the bright β Cephei stars. They performed high resolution, high signal-to-noise time-resolved spectroscopic observations of KP Per, V986 Oph and β Sco A with an aim of doing seismic modelling of these stars. Their results show that the HeI 6678 Å line profiles of KP Per and V986 Oph show variations (see Fig. 1.22). Stateva et al. (2010)

reported that these results will help them to verify three modes suggested for KP Per and two for V986 Oph. For β Sco A two independent verifications: (i) its short pulsation period and (ii) eclipses, would imply that this star is a member of binary system.

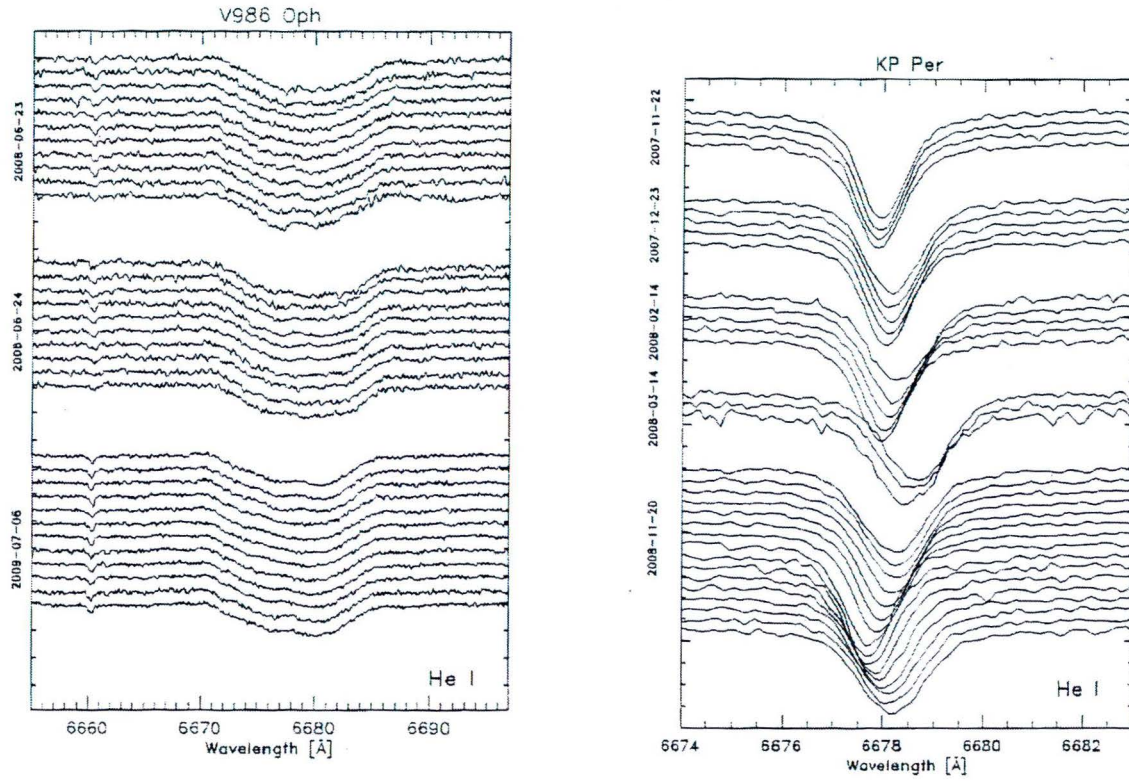


Figure 1.22: Figures showing HeI 6678 Å line profile variations of V986 Oph (left) and of KP Per (right) (taken from Stateva et al. 2010).

1.5 Slowly Pulsating B (SPB) stars

These are population I B-type stars showing multi-periodic light variations. SPB stars were discovered by Waelkens (1991) and they have spectral types ranging from B2 to B9 (Waelkens 1991). On the other hand Aerts, Christensen-Dalsgaard & Kurtz (2009) reported the spectral type range of SPB stars to be from B3 to B9, which draws a clear distinction from the β Cephei stars. The light variations in SPB stars have the periods within the range from 0.5 days to 3 days and the masses ranging from $2 M_{\odot}$ to $7 M_{\odot}$ (Aerts, Christensen-Dalsgaard & Kurtz 2009). The pulsation period range also sets SPB stars apart from β Cephei stars because it is longer for SPB stars as compared to that of β Cephei stars. The instability strip of SPB stars in the HR-diagram is located along the main-sequence, just below that of β Cephei stars (see Fig. 1.1), however, Miglio et al. (2007) mentioned the possibility of an overlap of the instability strip of SPB stars with that of β Cephei stars. This suggests that there are hybrid SPB- β Cephei stars. The pulsations in SPB stars were found to be driven by the κ mechanism (Dziembowski et al. 1993). The spectroscopic study showed the variations of the line profile of SPB stars. These variations, together with light variations are due to non-radial pulsations of high order g-modes. SPB stars are slowly rotators with rotational velocity ($v \sin i$) ≤ 100 km/sec and radial velocity variation amplitude less than 15 km/sec (Aerts & De Cat 2003).

The presence of SPB stars in the LMC was reported by Kolaczowski et al. 2004. The period of light variations of these first extragalactic SPB stars was found to be longer than 0.5 days. In addition, the hybrid SPB- β Cephei stars were also discovered in the LMC. The discovery of SPB stars in LMC was done using the OGLE II photometric data which was supplemented by MACHO photometric data. Fig. 1.23 shows the frequency spectra of one of the LMC SPB stars.

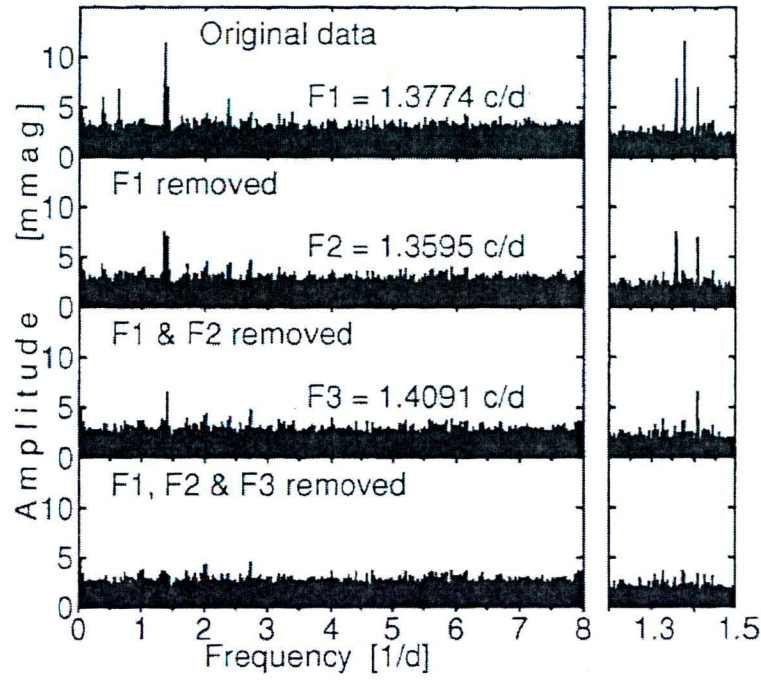


Figure 1.23: A figure showing consecutive steps of prewhitened Fourier periodograms of the multi-periodic SPB star, SC1-223227 in the LMC (taken from Kolaczowski et al. 2004).

The frequency spectra in Fig. 1.23 are given in the order of the removed detected frequencies (prewhitening). The SPB stars were also discovered in the open cluster NGC 371 of the SMC. Karoff et al. (2008) found 29 short-period B stars which are believed to be SPB stars.

1.6 Pulsating Be stars

Be stars are the type of B stars which are members of population I stars, with or without which have shown the Balmer line emission in their spectrum. The Balmer line emission is clearly the main feature distinguishing Be stars from β Cephei stars and other pulsating B stars, and it is believed to be due to the presence of a circumstellar equatorial disk ejected from the star itself (Aerts, Christensen-Dalsgaard & Kurtz 2009). There are various explanations for the presence of the disk around Be stars and Aerts, Christensen-Dalsgaard & Kurtz (2009) reported that since most Be stars are members of close binary systems, then the disk formation is due to the mass transfer from the companion. However, for the single Be stars, the disk formation is explained to be due to the critical rotation speed that these stars can attain. But, Neiner & Hubert (2008) argued that even though Be stars can rotate fast (with typical rotation velocity of 250 km/sec), their rotation velocity is not high enough to reach critical limit for matter to be ejected from the star and form the disk. It is further reported that Be stars show the presence of winds from the poles (polar winds). There are other suggestions of trying to explain the presence of the disk around Be stars, these include (i) magnetic field, which Neiner & Hubert (2008) mentioned that it could force stellar matter to flow along the field lines. This implies that the star will lose matter from the poles and this matter from both poles will collide around the equatorial region and the rotation of the star will drag it to form a disk. (ii) The beating of pulsation modes, this is described to be a possible angular momentum enhancement parameter that would add-up with the fast rotation of the star to reach the critical rotation limit. As a result the stellar matter will escape to form the disk around the equatorial stellar region.

Be stars were found to show light variability on different time scales, this was reported by Balona (1995) to be from 0.5 days to 2 days. This variation period is reported differently by Neiner & Hubert (2008) as being from 0.3 days to 2 days. The two kinds of light variations were reported to be present on Be stars, these are (i) variations due to rotation, pulsations, magnetic fields and polar winds. These variations are of the

order of days, (ii) there are also variations due to the changes in the circumstellar disk structure and sudden outbursts. These are in the order of months to decades. Neiner & Hubert (2008) noted that the amplitude of the line profile variations in Be stars is inversely proportional to the rotation velocity. Be stars are located on the same region as β Cephei and SPB stars in the HR-diagram, this was enough to convince Neiner & Hubert (2008) that the pulsations in these stars are driven by the κ mechanism.

Diago et al. (2009) did the analysis of photometric data of the CoRoT mission star, HD 50209. This is a B8 IVe star. The idea was to check if this star has the characteristics of Be stars as its location in HR-diagram suggested. It was found that HD 50209 is a multi-periodic Be star with four frequencies corresponding to g-modes with azimuthal order: $m = 0, -1, -2$ and -3 . The rotational period of this star was also found using the frequency of 0.679 cycles per day (Diago et al. 2009).

It was also discovered that there are extragalactic Be stars, from MACHO photometric time series. Diago et al. (2008) searched for short-term periodic variability in a sample of 313 B-type stars of the SMC. They found that from their sample, 76 stars showed the characteristics of Be stars.

Even though the reason behind the pulsations in β Cephei stars, Be stars and SPB stars was not understood for a long time, it was eventually discovered in the early 1990's that the classical κ mechanism is the mechanism driving pulsations in these stars. The following section gives a description of κ mechanism and its discovery.

1.7 The mechanism that drives pulsations in β Cephei stars

The mechanism that drives pulsations in β Cephei stars is called the κ mechanism. It is facilitated by ionization of metals. The κ mechanism in β Cephei stars gets triggered at the temperature of about 2×10^5 K (Pigulski & Kolaczowski 2002). At this temperature, a large number of electrons in iron-group atoms make bound-bound transitions, forming the iron opacity bump or Z bump (Carroll & Ostlie 2007). These transitions cause a small change in the density of atoms at the Z bump.

The κ mechanism is basically activated when the opacity at the Z bump region increases. The opacity increase is mostly caused by a small perturbation in the density of the iron-group atoms. When the star becomes opaque at Z bump region, the absorption of energy becomes more efficient, this energy absorption results in the increase in the temperature and eventually in the increase in gas pressure. For a star in hydrostatic equilibrium, the gas pressure is balanced by a gravitational force and these two quantities are related by the hydrostatic equilibrium equation:

$$\frac{dP}{dr} = -g\rho \quad (1.2)$$

where P is the gas pressure, g is gravity and ρ is the density of stellar materials being held together by pressure gradient force and gravitational force. However, the increased pressure at the Z bump will overcome the gravitational force and the stellar layers above the Z bump will expand. As the expansion occurs, the temperature of the layers will drop because of the decrease in density due to expansion. The temperature-density relation is the result of adiabatically expanding given as:

$$PV^\gamma = K \quad (1.3)$$

where P is pressure, V is the volume, K is the constant and $\gamma = \frac{n+1}{n}$ where n is called polytropic index. If the volume of the expanding star is related to its density by

$$V = \frac{1}{\rho} \quad (1.4)$$

where V is the volume and ρ is density, then the gas pressure becomes

$$P = K\rho^\gamma \quad (1.5)$$

Assuming that the gas causing expansion is perfect (ideal gas), pressure can be expressed as

$$P = \frac{\rho k_B T}{\mu m_u} \quad (1.6)$$

where k_B is Boltzmann's constant, T is the temperature, μ is a mean molecular weight and m_u is mass of a gas. Therefore,

$$T \propto \rho^{\gamma-1} \quad (1.7)$$

and this will result in a decrease in pressure. As a result, gravitational force will dominate pressure gradient force because the pressure will decrease such that expansion overshoots the equilibrium point, and thus the layers will contract. Contraction will cause the temperature to increase and so will the pressure, resulting in expansion again. This contraction and expansion of stellar layers cause the variability of the light output of the star.

The pulsations driven by the κ mechanism in stars can be thought of as the piston in a car's engine. The piston is the suitable analogy because when the connection rod is fully pulled out, the gas inside the cylinder, in-front of the piston is not under pressure. When the connection rod is pushed inwards, the gas in-front of the connection rod heats up and that causes the gas pressure to increase. As a result the gas pressure will force the connection rod outwards. The inwards and outwards motion of the connection rod resembles the expansion and contraction of the stellar layers when the star is pulsating due to opacity increase. This kind of pulsation is called radial pulsation, and is represented by spherical harmonic degree of zero (i.e. $l = 0$). There are many other pulsations with higher l values and the star with most known l values is the Sun.

Several attempts were made to find pulsational instability in stellar models of β Cephei stars, but it was following the availability of the new opacity tables called OPAL by Lawrence Livermore National Laboratory that Cox et al. (1992); Moskalik & Dziembowski (1992) and Kiriakidis et al. (1992) independently found that β Cephei stars are pulsationally unstable for the radial fundamental mode and for low-degree non-radial modes (Osaki 1993).

Kiriakidis et al. (1992) investigated the evolution and stability of stars in the mass range from $9 M_{\odot}$ to $20 M_{\odot}$ with initial metallicities of $Z = 0.02$ and $Z = 0.03$. They found unstable pulsations for metallicity of $Z = 0.03$ and mass of $15 M_{\odot}$ in the temperature range of $4.33 < \log T_{\text{eff}} < 4.43$.

Moskalik & Dziembowski (1992) conducted a linear stability survey for a number of sequences of unfitted envelope models in the parameter range corresponding to β Cephei stars. Their calculations were done using new opacities mentioned above and their composition parameters were $X = 0.7$, $Z = 0.03$ (some of their calculations were done using $Z = 0.04$) and Anders-Grevesse mixture of heavy elements. They found an unstable fundamental radial mode and claimed that such instability "*should persist in non-radial modes of similar and lower frequencies*". Moskalik & Dziembowski (1992) concluded that the instability they found was driven by the κ mechanism where there is a bump in opacity at temperatures near 2×10^5 K. The cause of the instability was reported to be due to a slight surplus of the driving where the damping occurs over the entire stellar interior. The agreement between the theoretical instability strip calculated by Moskalik & Dziembowski (1992) and observational data was found and is shown in Fig. 1.24.

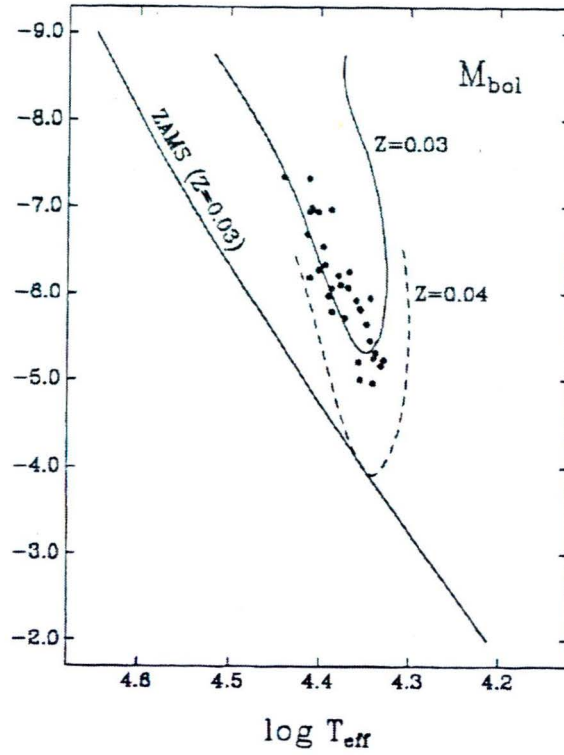


Figure 1.24: A figure showing the HR diagram of β Cephei stars (dots) and theoretical instability strips of the radial fundamental mode. The diagonal straight line represents the Zero Age Main Sequence (ZAMS) for a metallicity of 0.03, the solid line curve is the theoretical instability strip for a metallicity of 0.03 and the dashed curve is the theoretical instability strip for a metallicity of 0.04 (taken from Moskalik & Dziembowski 1992).

In Fig. 1.24 the two instability strips are shown for $Z = 0.03$ (solid curve) and for $Z = 0.04$ (dashed curve). As it can be seen, it was found that the width of these strips is sensitive to metallicity, i.e. the increase in metallicity produces an instability strip with larger width. However the sparsity of opacity tables causes uncertainties which prevented Moskalik & Dziembowski (1992) from quantifying metallicity-instability strip width dependence.

Miglio et al. (2007) investigated the effect of uncertainties in the metal mixture on the instability strips of B stars, and also on the frequency domain of excited modes. Their analysis was done using two opacity tables; OPAL (Rogers & Iglesias 1992) and

OP (Opacity Project and Seaton 1996), and two mixtures of different metals; GN93 mixture (Grevesse & Noels 1993) and AGS05 + Ne mixture (Asplund et al. 2005). The main-sequence models with masses ranging between $2.5 M_{\odot}$ and $12 M_{\odot}$ were used on the computations. When comparing the computations done with OPAL opacities, GN93 metal mixture and OP opacities, Miglio et al. (2007) found that OP opacities computations brought the suggestion of the excited high-order g-modes in hotter stars. On considering the AGS05 + Ne metal mixture, it was found that higher overtone modes were also excited. They also found that even for metallicities as low as $Z = 0.01$ there were excited modes, particularly in β Cephei pulsations. This analysis was motivated by the discovery of β Cephei stars in low-metallicity environment, i.e. Magellanic Clouds (see Pigulski & Kolaczowski 2002). They believed that there could be an underestimation on the "*iron opacity bump*" from stellar models.

1.8 A description of the OGLE database

The OGLE II database was used for selecting the fields/targets observed in this thesis, this section gives a brief description of OGLE II and the description of other OGLE missions.

The Optical Gravitational Lensing Experiment (OGLE¹) is a long term observational project which is meant to detect dark matter through microlensing phenomenon. By the time of writing this thesis, there were three completed observation campaigns, OGLE I, OGLE II and OGLE III campaigns. The first campaign (OGLE I) started in 1992 with the construction of a 1.0-m telescope at Chile, Las Campanas Observatory. This telescope, equipped with a 2048×2048 Ford/Loral CCD camera, saw its first light in 1996 and normal observations started a year later, in 1997. This first campaign found nineteen microlensing phenomena towards a Galactic bulge. The second campaign started in early 1997 using a 1.3-m Warsaw telescope which is also located at Las Campanas Observatory. The Warsaw telescope was equipped with 2048×2048 SITe CCD camera and its properties are shown on Table 1.2.

Pixel size	24 μ m
Scale	0.417 arcsec/pixel
Gain (medium mode)	7.1e ⁻ /ADU
Readout noise (medium mode)	6.3e ⁻
Gain (slow mode)	3.8e ⁻ /ADU
Readout noise (slow mode)	5.4e ⁻

Table 1.2: A table showing the properties of the SITe CCD camera used during the OGLE II observation campaign.

¹see <http://ogle.astrouw.edu.pl/>

The results from the OGLE II observations, which took four years to be completed, are shown on Table 1.3. The OGLE II campaign observed the stars in a Galactic bulge, stars in the Large Magellanic Cloud (LMC) and stars in the Small Magellanic Cloud (SMC).

Targeted object	Number of fields	Sky coverage (sq.deg)	Number of stars observed	Number of measurements
Galactic bulge	49	11	30.5×10^6	9.4×10^9
LMC	26	4.5	6.8×10^6	2.7×10^9
SMC	11	2.4	2.2×10^6	0.7×10^9

Table 1.3: A table showing the results from the OGLE II observation campaign.

By mid-2002, the OGLE II campaign had detected about three hundred and fifty microlensing phenomena and in 2003 this number went up by hundred and ten phenomena. The third observation campaign, OGLE III, started during mid-2001 and was also carried out from Las Campanas Observatory, using a 1.3-m Warsaw telescope. For this campaign, the new “second generation” CCD mosaic camera, with the properties shown on Table 1.4, was used. Each mosaic chip was a SITe ST-002a CCD detector.

CCD mosaic size	8192×8192 pixels
CCD detector size	2048×4096 pixels
pixel size	15 μm
Scale	0.26 arcsec/pixel
Mosaic’s full field of view	35×35 arcmins

Table 1.4: A table showing the properties of the “second generation” CCD mosaic camera used on the OGLE III observation campaign.

The OGLE mission divided the LMC into 26 rectangular regions, our observations described in chapter 2 were done in two of those twenty-seven regions. Fig. 1.25 shows the OGLE regions on the LMC. It can be seen that the regions 21, 22, 23, 24, 25 and 26 were made on the external clusters of LMC whereas all others were along the bar of the LMC.

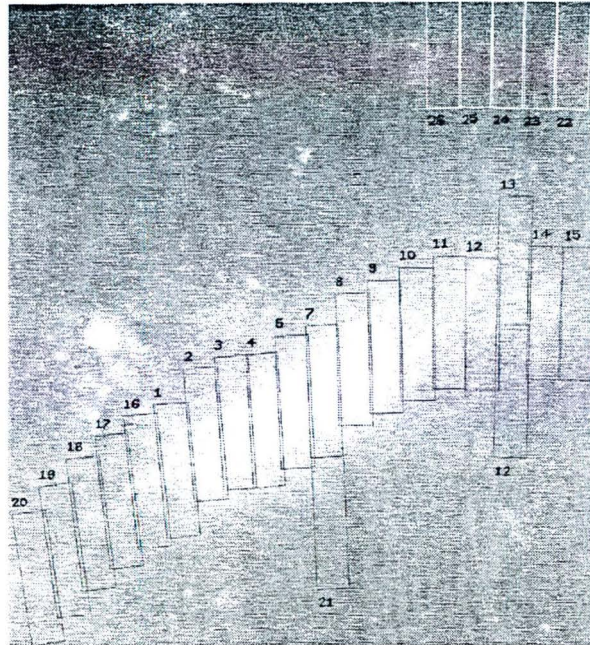


Figure 1.25: The figure showing the LMC and twenty-seven rectangular regions within which OGLE observations were made (see <http://ogledb.astrouw.edu.pl/ogle/photdb/>).

Even though the OGLE project was basically meant for dark matter detections, the data collected during each campaign were also useful for other astronomical purposes. One such purpose was done by Pigulski and Kolaczowski in 2002 to search and discover the first extra-Galactic β Cephei stars in LMC, and that discovery is the motivation behind the aim of this thesis.

Chapter 2

Target Selection and Observations

This chapter first describes how the candidate β Cephei stars we wanted to observe were selected from the OGLE database. The parameters used to extract the data from the OGLE II database are given, then the coordinates, V magnitude and colours (V-I and B-V) of the selected candidates are presented. The details of observations of selected candidates are also presented.

2.1 Target Selection

In order to do observations of β Cephei stars in the LMC, such targets had to be selected carefully. The targets were selected using parameters in Table 2.1. The OGLE II database was chosen because of the discovery of the first extra-Galactic β Cephei stars in the LMC from this second mission of OGLE (see Pigulski & Kolaczowski 2002).

Filter	Magnitude range
V	14.0 to 16.5
V-I	-2 to 0
B-V	-2 to 0

Table 2.1: A list of parameters used to extract the time series data from the OGLE II database for the selection of β Cephei stars.

The selection parameters on Table 2.1 were chosen such that they are consistent and close to those used by Pigulski & Kolaczowski (2002). Before they could decide on the limiting magnitudes for their target selection, Pigulski & Kolaczowski (2002) considered the visual extinction effect on the brightness of the LMC's stars and to account for this they added 1 magnitude to their limiting magnitude, and the values: $V < 18$ magnitude and $V-I < 0.5$ magnitude were used as limiting parameters. In our case, we also had to consider the capability of the telescope to be used for observations and hence our faint end was 16.5 magnitude in the V-filter. The parameters on Table 2.1 were entered on the OGLE II data extraction query page, such query page is shown in Fig. 2.1 with the parameters entered. As a result 4738 candidate stars were extracted and their light curves as well. The light curves were Fourier transformed using the program called *olev.f* provided by Dr. Balona.

OGLE BVI Maps Query Page

Select OGLE target:

☐ Galactic Bulge ☒ LMC ☐ SMC

Enter values or ranges of parameters, check appropriate Use boxes (Uncheck Query box below):

Show	Sort	Parameter	Use	Value/Range	Description
☐	☐	Field	☐		OGLE field name
☐	☐	StarID	☐		Star no. in field catalog
☒	☐	RA	☐		Right Ascension (J2000) ¹
☒	☐	Decl	☐		Declination (J2000) ¹
☐	☐	X	☐		X pixel coord ¹
☐	☐	Y	☐		Y pixel coord ¹
☒	☐	V	☒	14.0 16.5	Mean V-band magnitude
☒	☐	V_I	☒	-2 0	Mean V-I color index
☐	☐	I	☐		Mean I-band magnitude
☒	☐	Vgood	☐		No. of V-band good points
☐	☐	Vbad	☐		No. of V-band bad points
☐	☐	Vsig	☐		V mag standard deviation
☐	☐	Igood	☐		No. of I-band good points
☐	☐	Ibad	☐		No. of I-band bad points
☐	☐	Isig	☐		I mag standard deviation
☒	☐	B_V	☒	-2 0	Mean B-V color index (LMC/SMC only)
☐	☐	B	☐		Mean B-band magnitude (LMC/SMC only)
☐	☐	Bgood	☐		No. of B-band good points (LMC/SMC only)
☐	☐	Bbad	☐		No. of B-band bad points (LMC/SMC only)
☐	☐	Bsig	☐		B mag standard deviation (LMC/SMC only)

¹ RA/Dec, X/Y may also specify a circle or rectangle centered on a point, see Query Help for details;
RA format: HH:MM:SS or H.HHHHH, Decl: $\pm$ DD:MM:SS or $\pm$ D.DDDDD

Enter SQL query using the above parameter names (Check Query box below):

☒ Query: `SELECT objects FROM db WHERE`
`V>=14 and V<=16.5 and V_I>=-2 and V_I<=0.0 and B_V>=-2 a`

Sort ☒ ascending ☐ descending ☐ Sexag. RA/Dec output

Check Show boxes above for the parameters to display, 500 objects per page, max of objects

Submit Query Note: Depending on the target and query it may take a while to complete.

Figure 2.1: A figure showing the OGLE Photometry Query Page where the parameters on Table 2.1 were entered to extract the data for target selection (see <http://ogledb.astrouw.edu.pl/ogle/photdb/>).

A python program (attached in section A.1 of the Appendix) was used to select the targets with the smallest False Alarm Probability (FAP) from *olev.f* output, and the cut-off for the FAP we used is 0.05. FAP is the probability of getting the value “S”, or the value greater than “S”, in the set of several measurements which are independent of each other. A similar description of FAP is given by Sturrock & Scargle (2010) as “*the probability that at least one out of M independent power values in a prescribed search band of a power spectrum, computed from a white-noise time series, is expected to be as large as or larger than a given value*”. Mathematically, FAP is denoted by $FAP(S|M)$, which means FAP of finding S in M independent measurements, in our case S stands for detected frequency signals on the frequency spectra and M stands for the magnitude measurements. The expression for calculating FAP is an exponential function presented by Sturrock & Scargle (2010) as follows:

$$FAP(S|M) = 1 - (1 - e^{-S})^M \quad (2.1)$$

Another selection criterion was the pulsation frequency, the candidates with frequency greater than 3 cycles per day (c/d) were selected as promising. From these promising candidates we searched for those close to each other. This was important for the observation since by having more than one target on one frame meant observing more than one target concurrently. However, none of the selected candidates were that close to each other.

The main selection criterion was to inspect the frequency spectrum of each candidate whose FAP was ≤ 0.05 and frequency ≥ 3 c/d. This inspection was mainly focussed on the peaks within the frequency range from 3 c/d to 15 c/d, which is the frequency range for β Cephei pulsations. Any target having a significant peak, which is high enough above the noise level within that frequency range was considered a potential β Cephei star to be included in the candidate list.

The noise level for each frequency spectrum was visually determined. We then visually inspected the Fourier spectra for aliases. An alias in the frequency spectrum is

caused by the gaps in the data due to the fact that during the daytime a star is not visible, and is not observed. From the power spectra of more than hundred candidates which satisfied criteria of $FAP \leq 0.05$ and frequency ≥ 3 c/d, sixteen candidates were found to have convincing peaks within a frequency range of β Cephei stars. To check if any of these sixteen candidate β Cephei stars have counterparts in the published literature, we used SIMBAD¹ coordinates query and found that only star lmc_sc7_i_207004 is located 0.16 arcsec from the known β Cephei star OGLE J051841.98-691051.9. All other candidate β Cephei stars were not published and not classified on SIMBAD. The light curves and frequency spectra of these sixteen candidates are shown on Figs. 2.2-2.17. The Nyquist frequencies of the spectra were determined with FAMIAS 1.01² (Zima 2008).

¹<http://simbad.cfa.harvard.edu/simbad/sim-fcoo>

²Nyquist frequencies were determined with the software package FAMIAS developed in the framework of the FP6 European Coordination Action HELAS (<http://www.helas-eu.org/>)

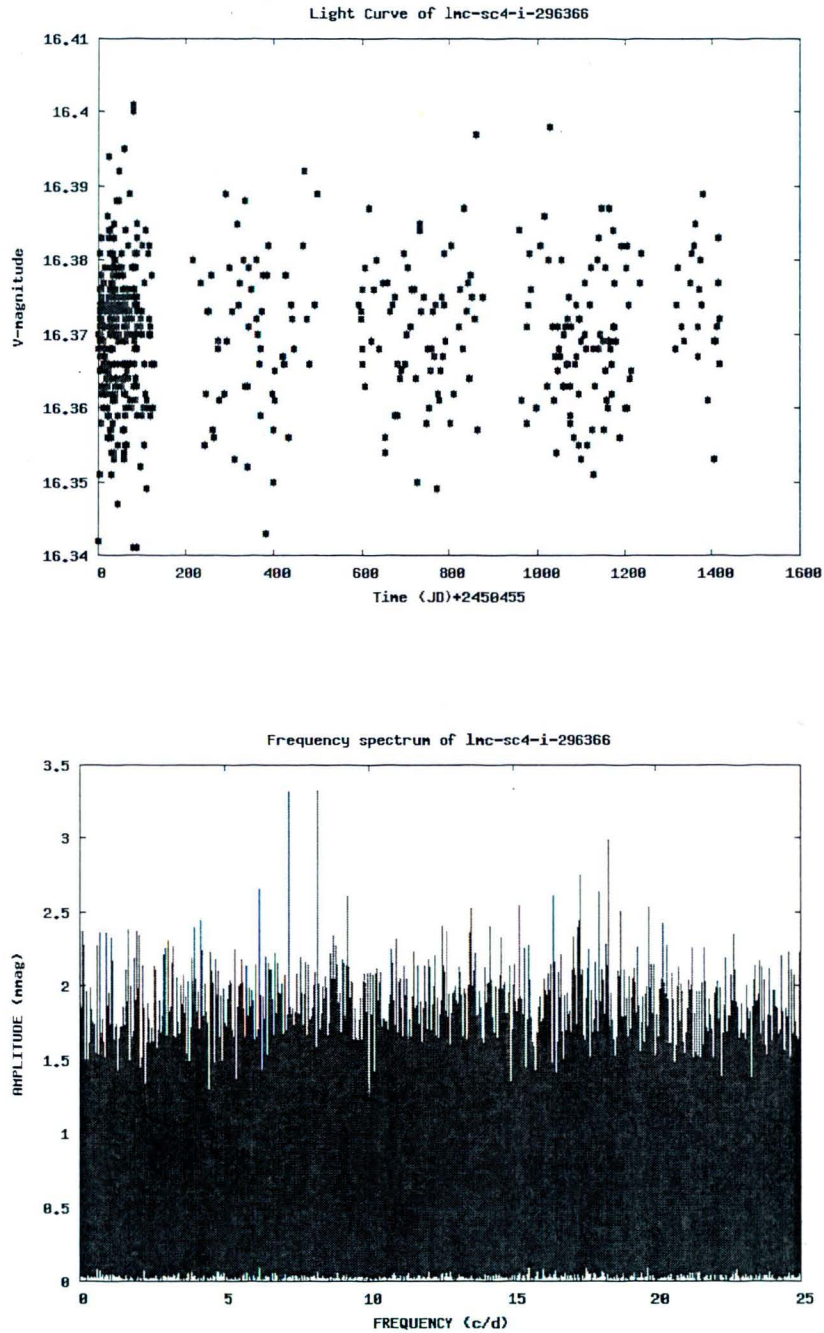


Figure 2.2: The top plot shows the OGLE light curve of `lmc_sc4_i_296366`, and the bottom plot is its frequency spectrum with a Nyquist frequency of 9.278 c/d. The highest peaks between 5 and 10 c/d look promising. Also notice a noise level which is less than 2.5 mmag on the spectrum.

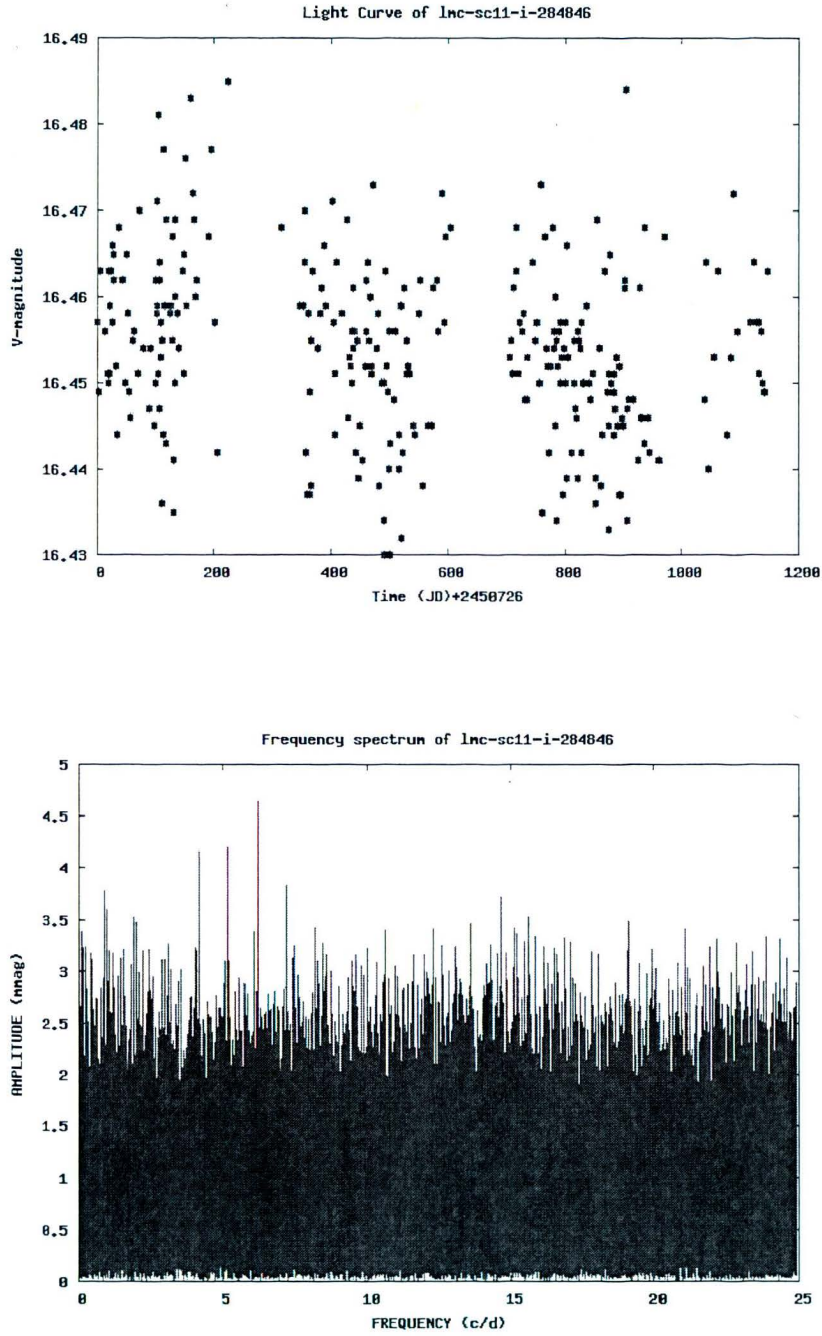


Figure 2.3: The top plot shows the OGLE light curve of `lmc_sc11_i_284846`, and the bottom plot is its frequency spectrum with a Nyquist frequency of 8.224 c/d. The highest peak just around 5 c/d looks promising and the other two promising peaks are between 0 and 5 c/d, even though they might be aliases given the noticeable gaps on the light curve.

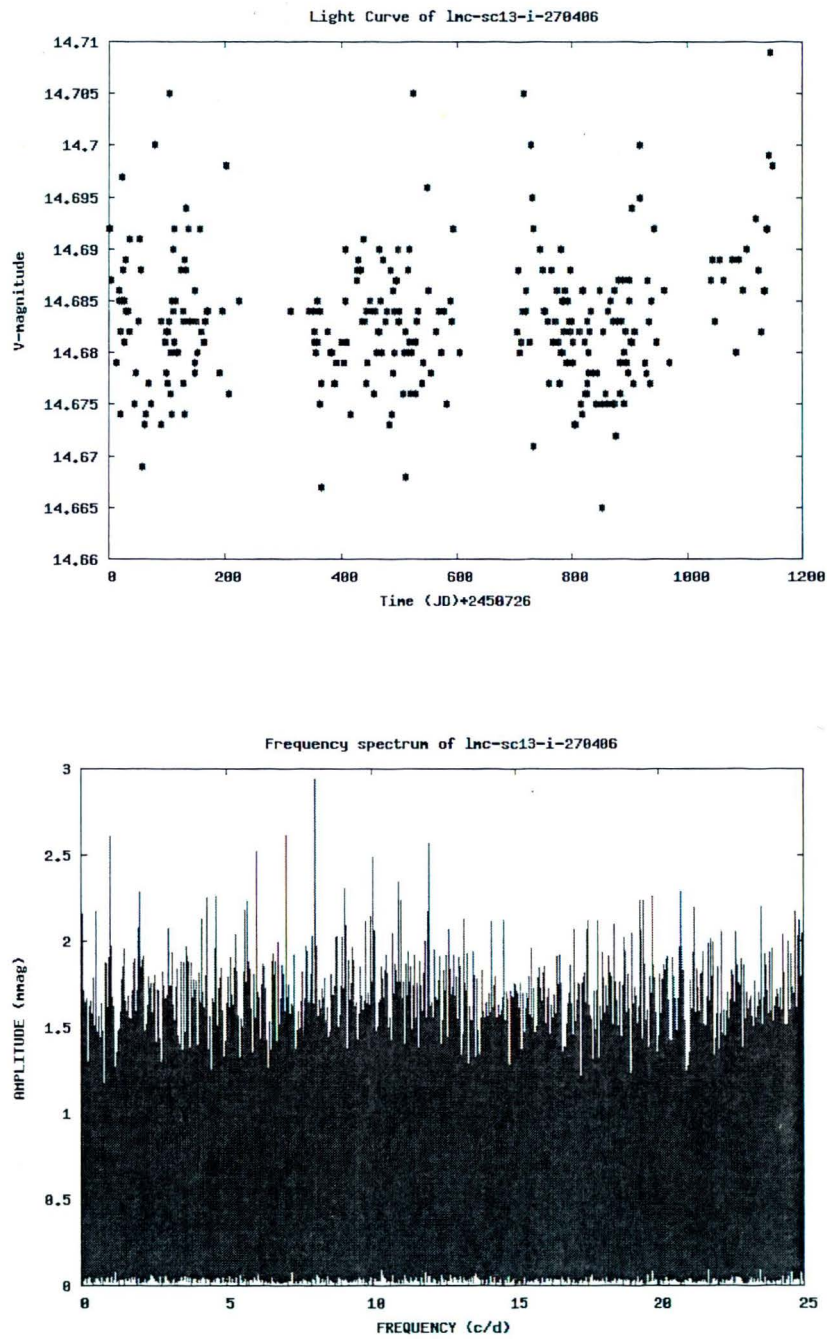


Figure 2.4: The top plot shows the OGLE light curve of lmc_sc13_i_270406, and the bottom plot is its frequency spectrum. The highest peak between 5 and 10 c/d looks promising.

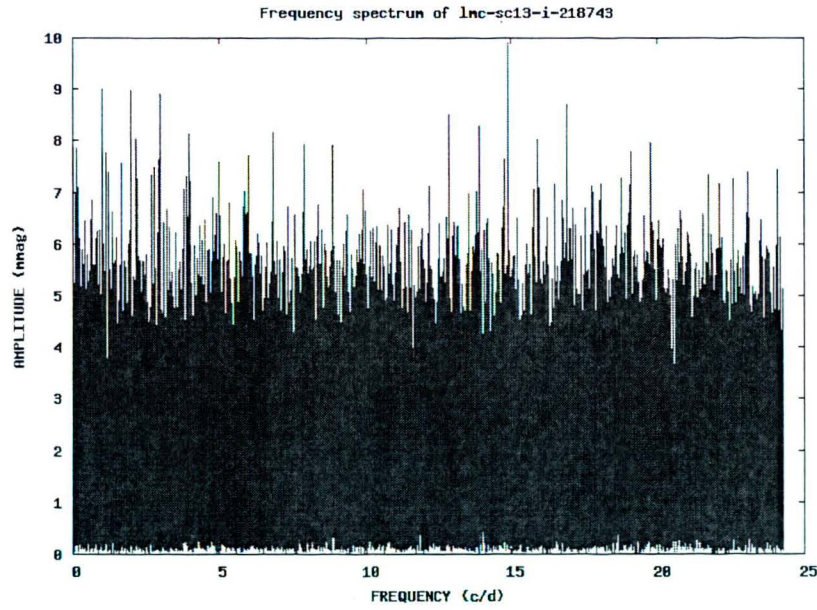
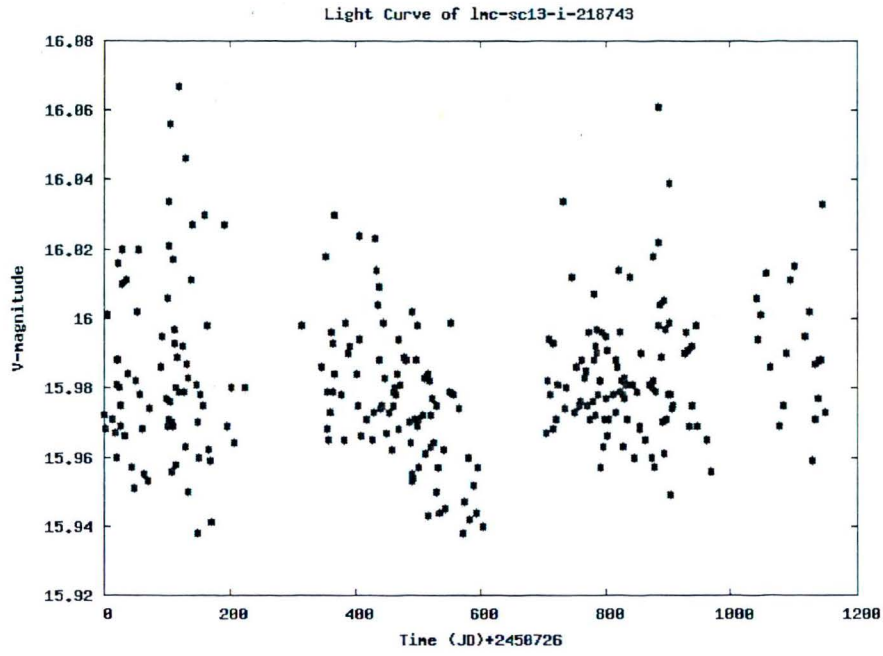


Figure 2.5: The top plot shows the OGLE light curve of lmc_sc13_i_218743, and the bottom plot is its frequency spectrum. The maximum peak between 10 and 15 c/d is promising. The other noise-outstanding peaks could be aliases as it can be seen from the light curve that there are gaps.

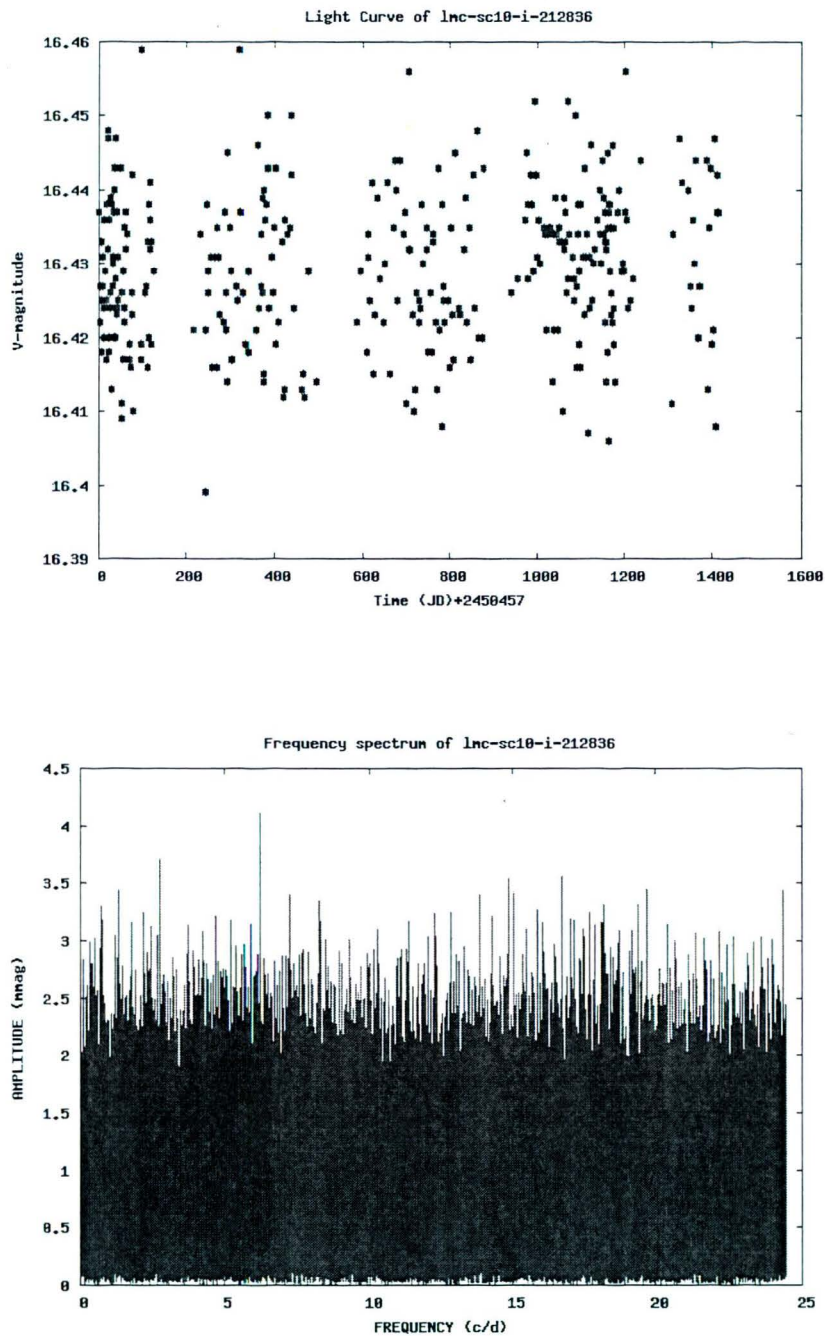


Figure 2.6: The top plot shows the OGLE light curve of lmc_sc10_i_212836, and the bottom plot is its frequency spectrum with a Nyquist frequency of 7.880 c/d. There are two peaks which are promising on this spectrum, the first is between 5 and 10 c/d and the second is between 0 and 5 c/d. The latter is however as high as the noise level.

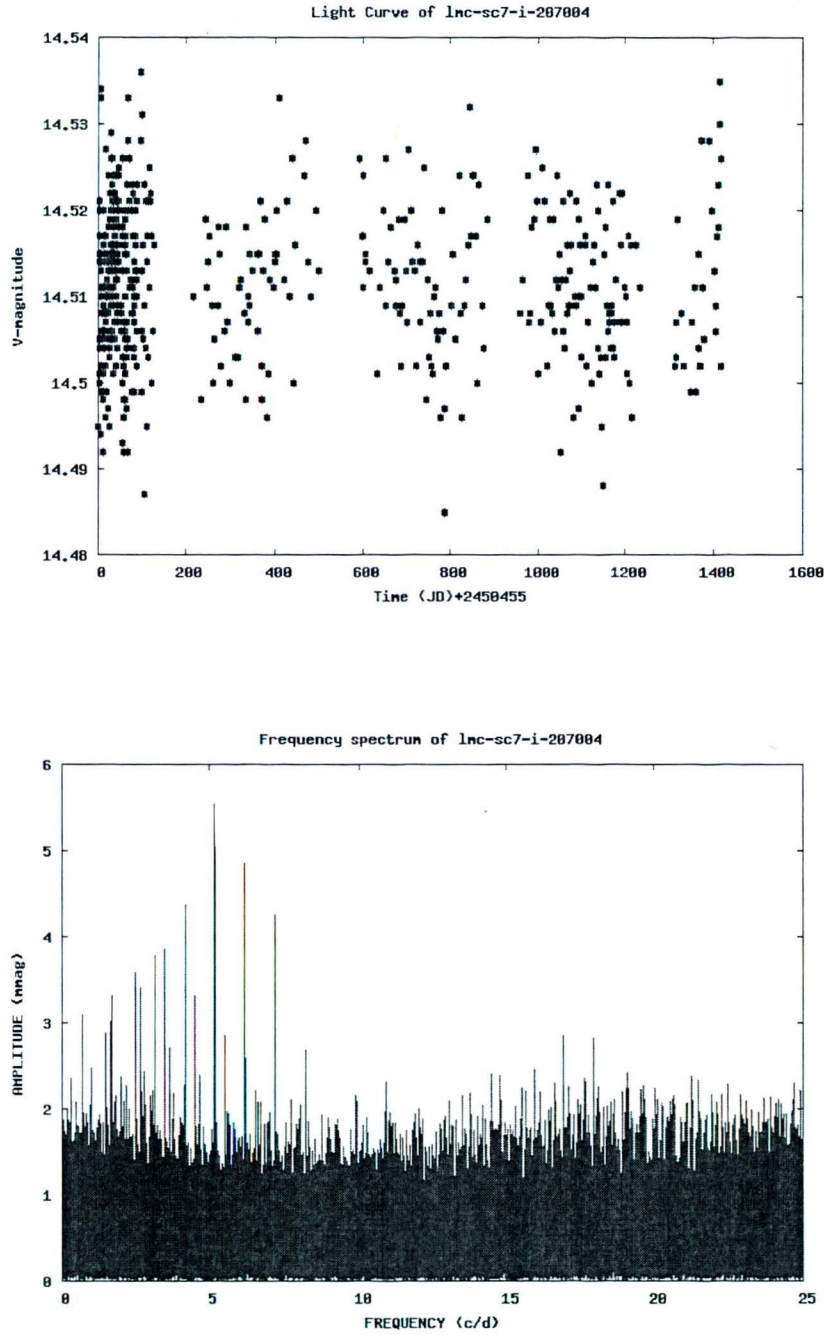


Figure 2.7: The top plot shows the OGLE light curve of `lmc_sc7_i_207004`, and the bottom plot is its frequency spectrum with a Nyquist frequency of 9.338 c/d. The maximum peak around 5 c/d is promising and also notice the noise level which is around 2 mmag on this spectrum. The other peaks might be aliases given the gaps on the light curve.

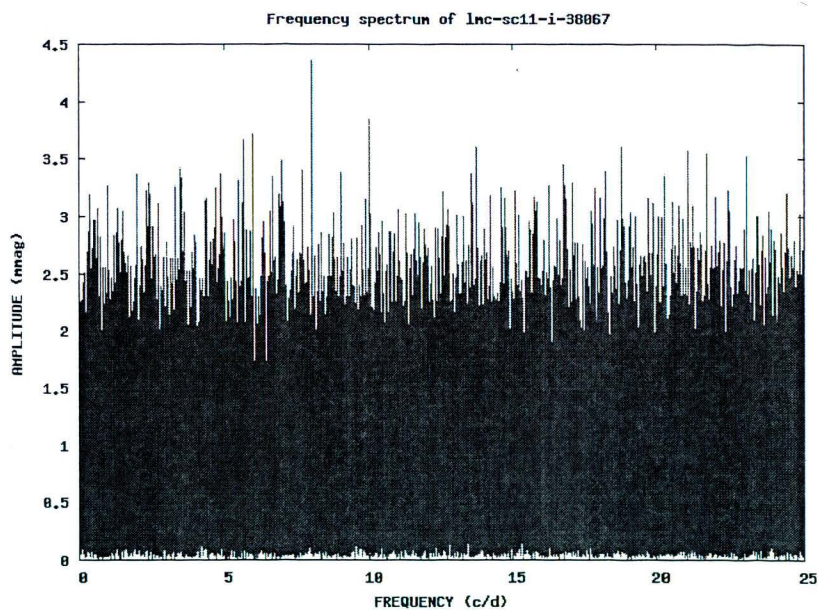
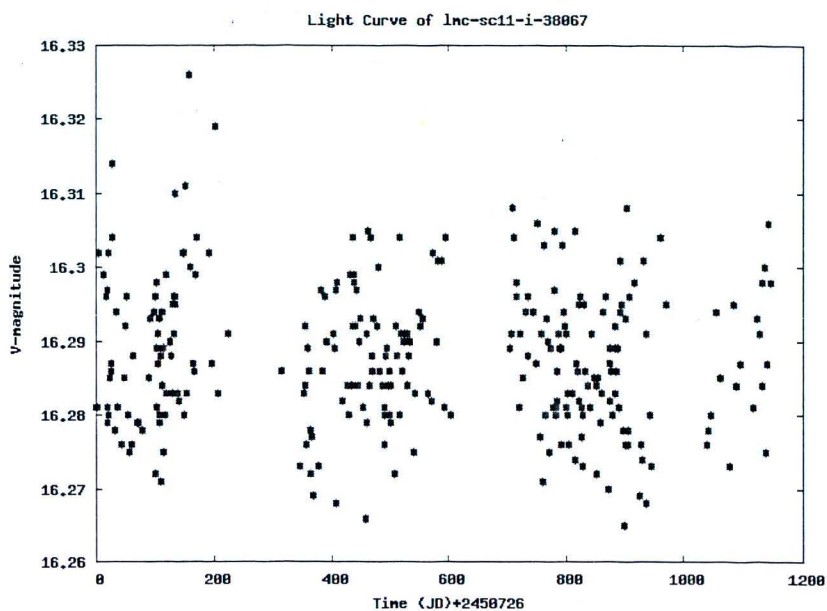


Figure 2.8: The top plot shows the OGLE light curve of lmc_sc11_i_38067, and the bottom plot is its frequency spectrum with a Nyquist frequency of 8.224 c/d. The highest peak between 5 and 10 c/d is promising.

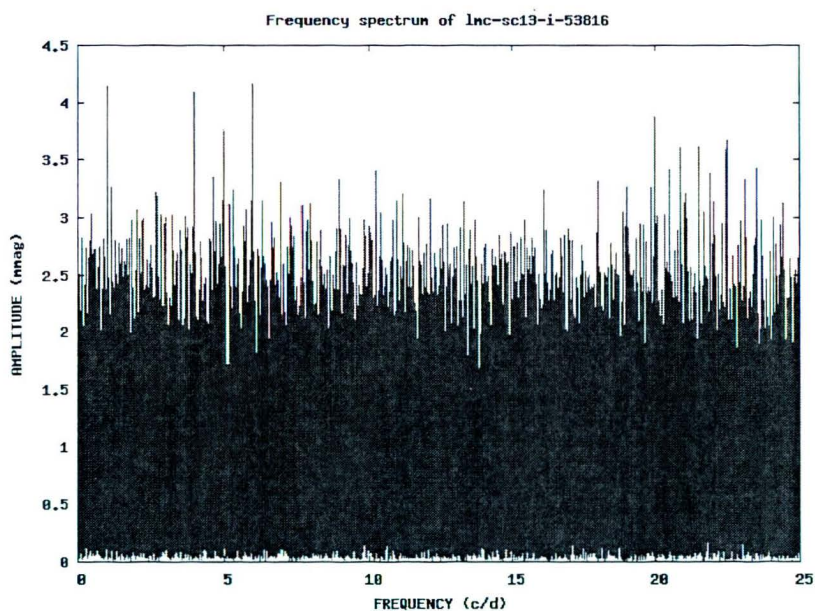
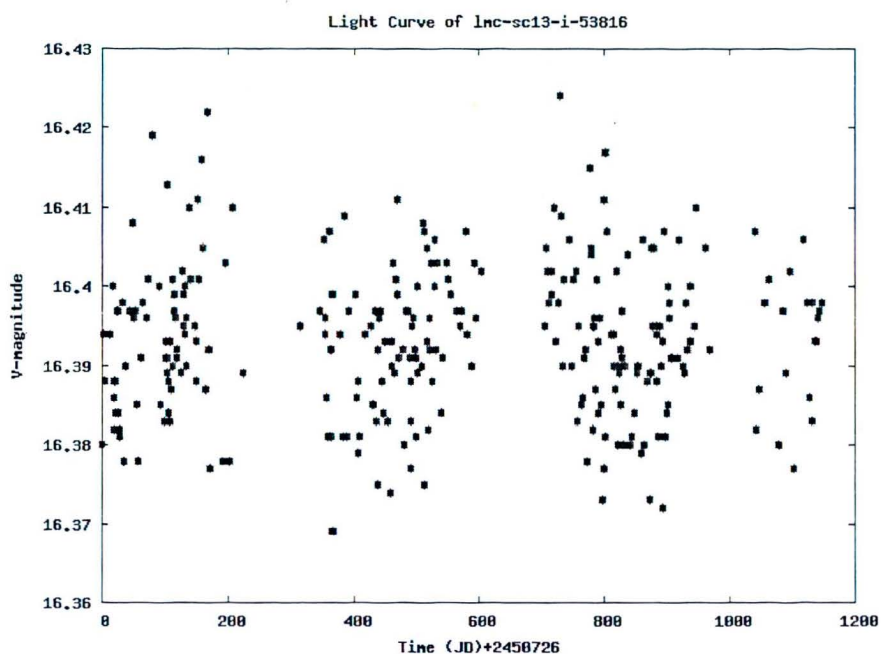


Figure 2.9: The top plot shows the OGLE light curve of lmc_sc13_i_53816, and the bottom plot is its frequency spectrum. There are three promising peaks between 0 and 5 c/d and also just above 5 c/d.

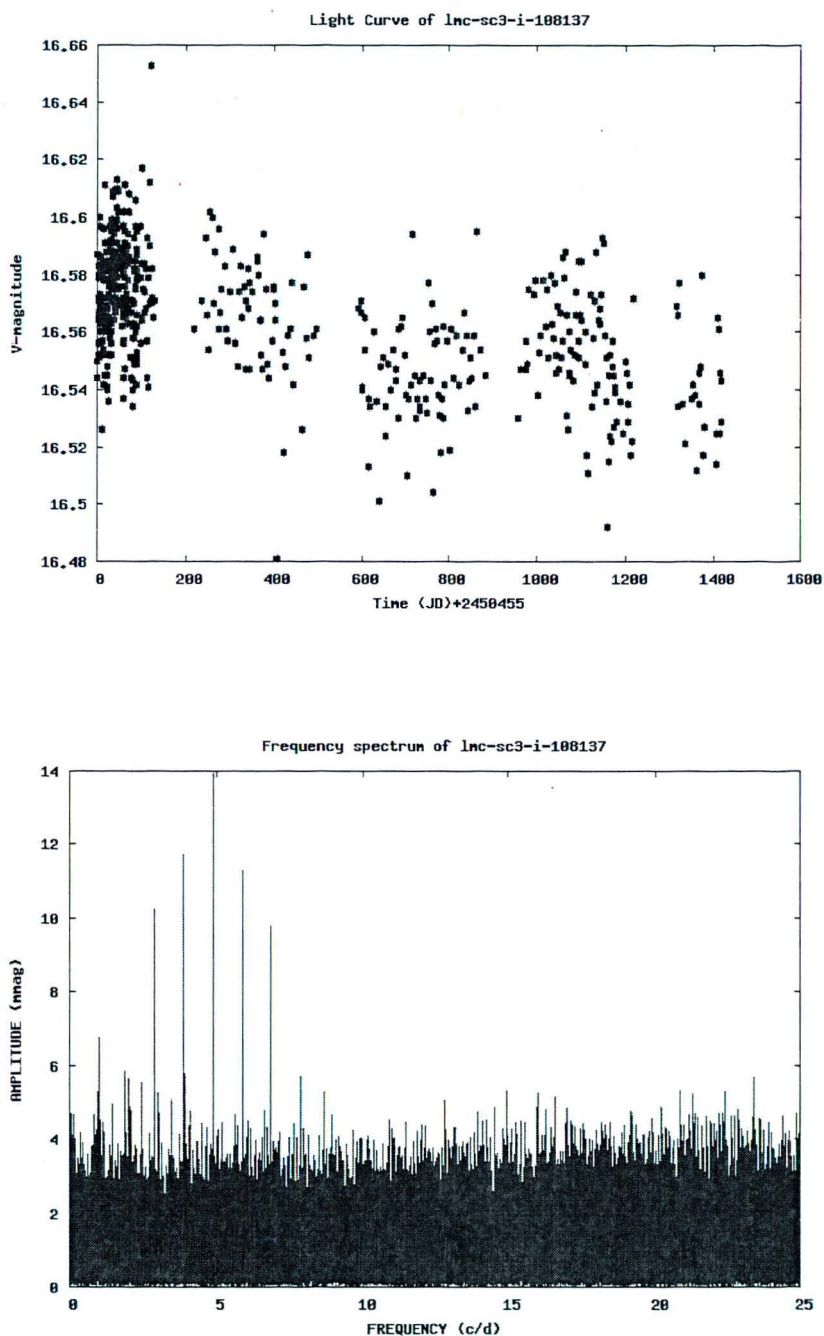


Figure 2.10: The top plot shows the OGLE light curve of lmc_sc3_i_108137, and the bottom plot is its frequency spectrum with a Nyquist frequency of 9.40 c/d. The highest peak between 0 and 5 c/d is promising and other peaks on both sides of the highest peak look more like aliases. Also notice the noise level on this spectrum which is around 5 mmag.

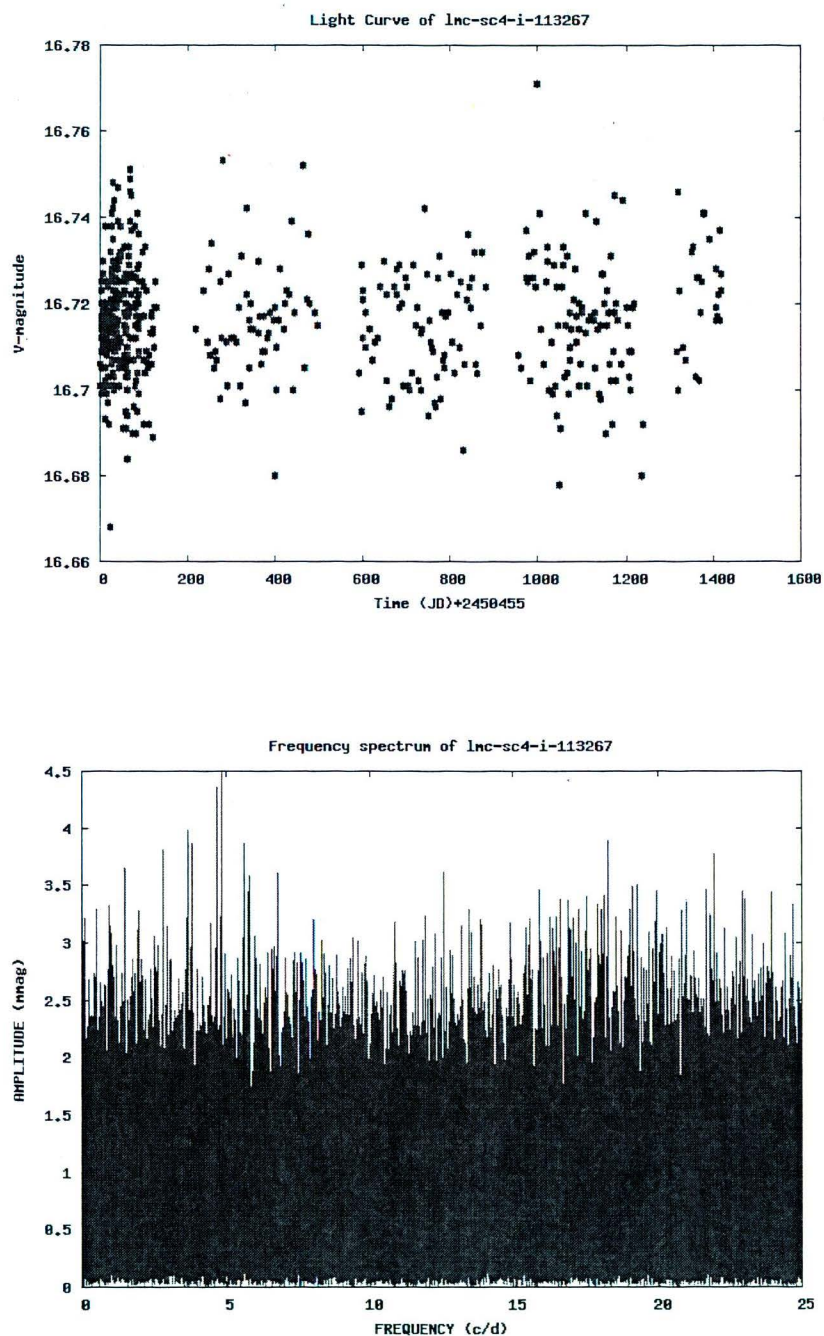


Figure 2.11: The top plot shows the OGLE light curve of lmc_sc4_i_113267, and the bottom plot is its frequency spectrum with a Nyquist frequency of 9.278 c/d. The highest peak between 0 and 5 c/d looks promising and the spectrum has noise level of about 3 mmag.

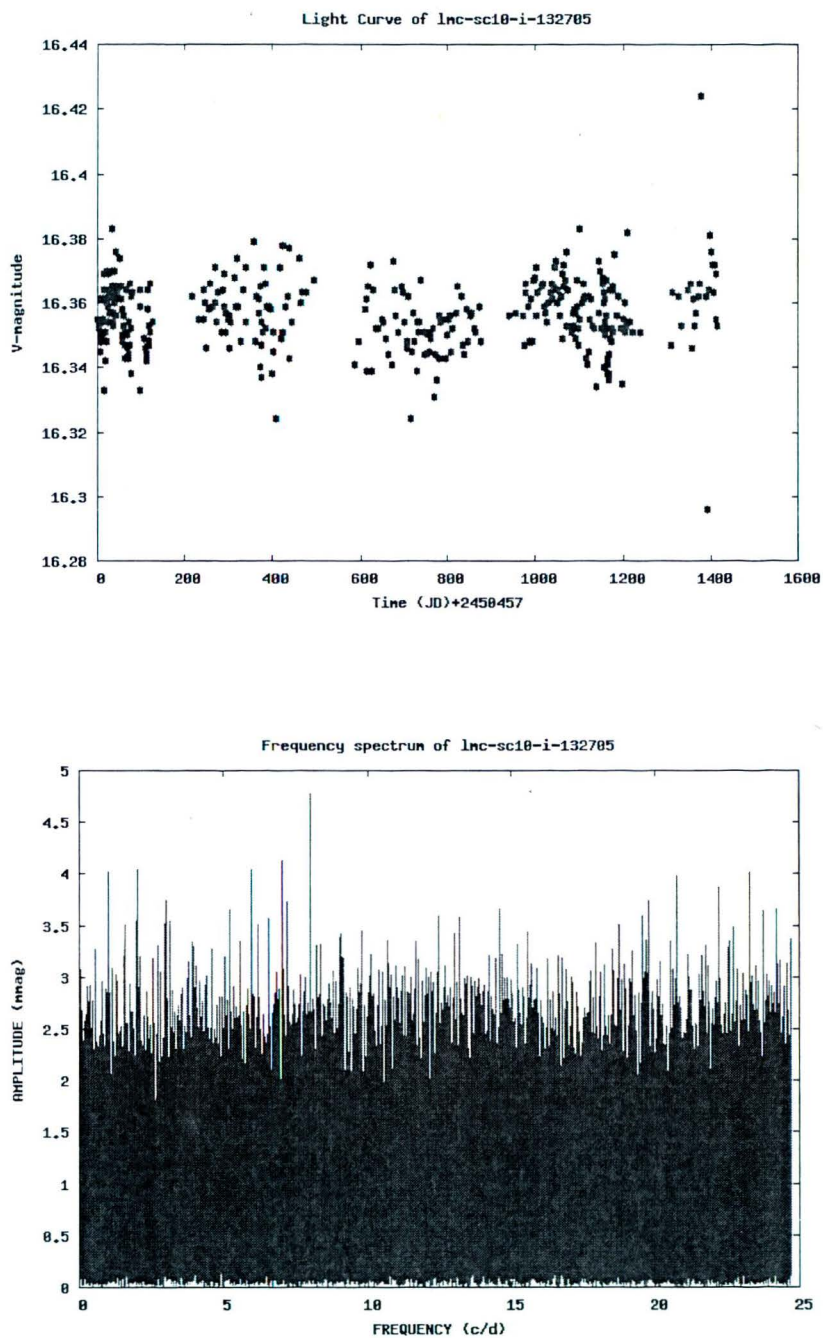


Figure 2.12: The top plot shows the OGLE light curve of lmc_sc10_i_132705, and the bottom plot is its frequency spectrum with a Nyquist frequency of 7.880 c/d. The highest peak between 5 and 10 c/d looks promising. The other peaks are hidden by the noise level of about 3 mmag on this spectrum.

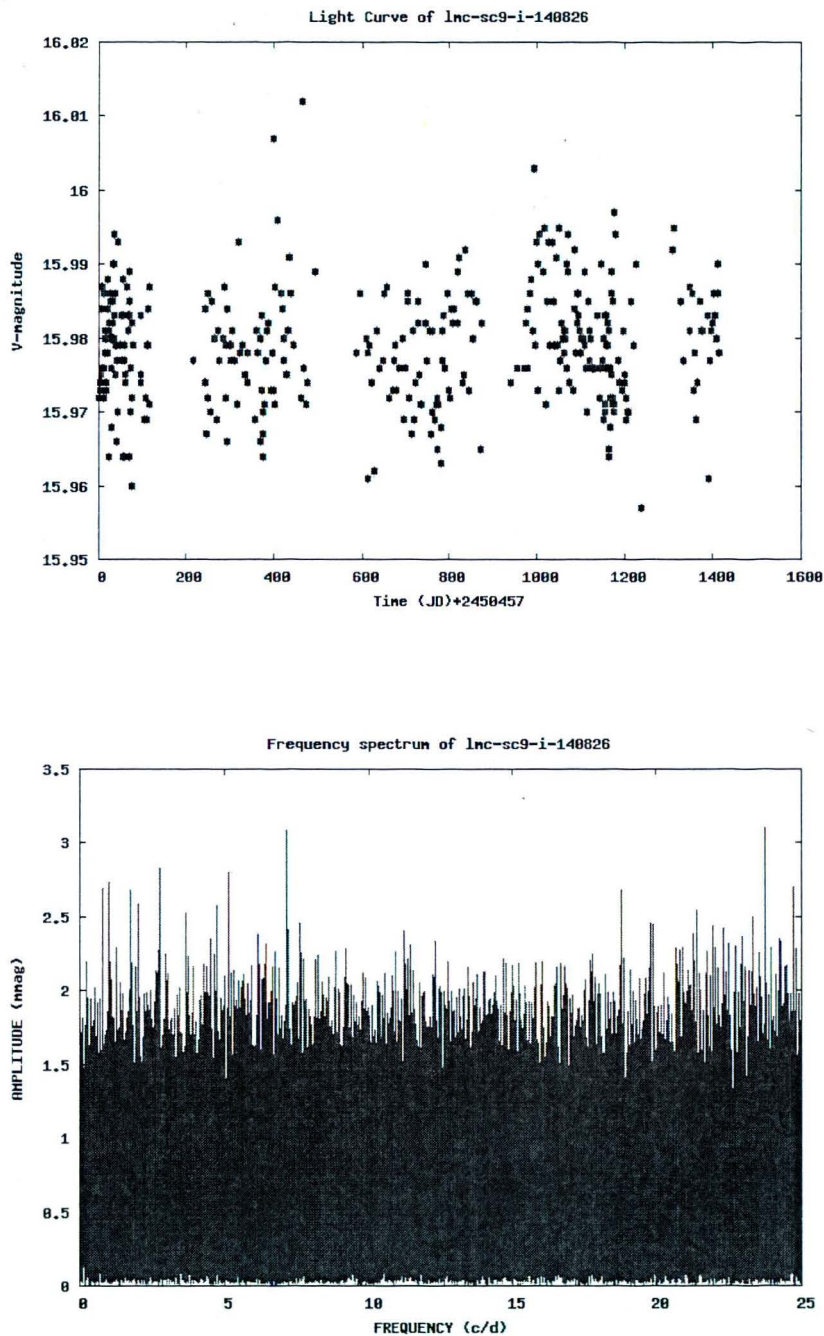


Figure 2.13: The top plot shows the OGLE light curve of `lmc_sc9_i_140826`, and the bottom plot is its frequency spectrum with a Nyquist frequency of 6.846 c/d. The peak between 5 and 10 c/d looks promising.

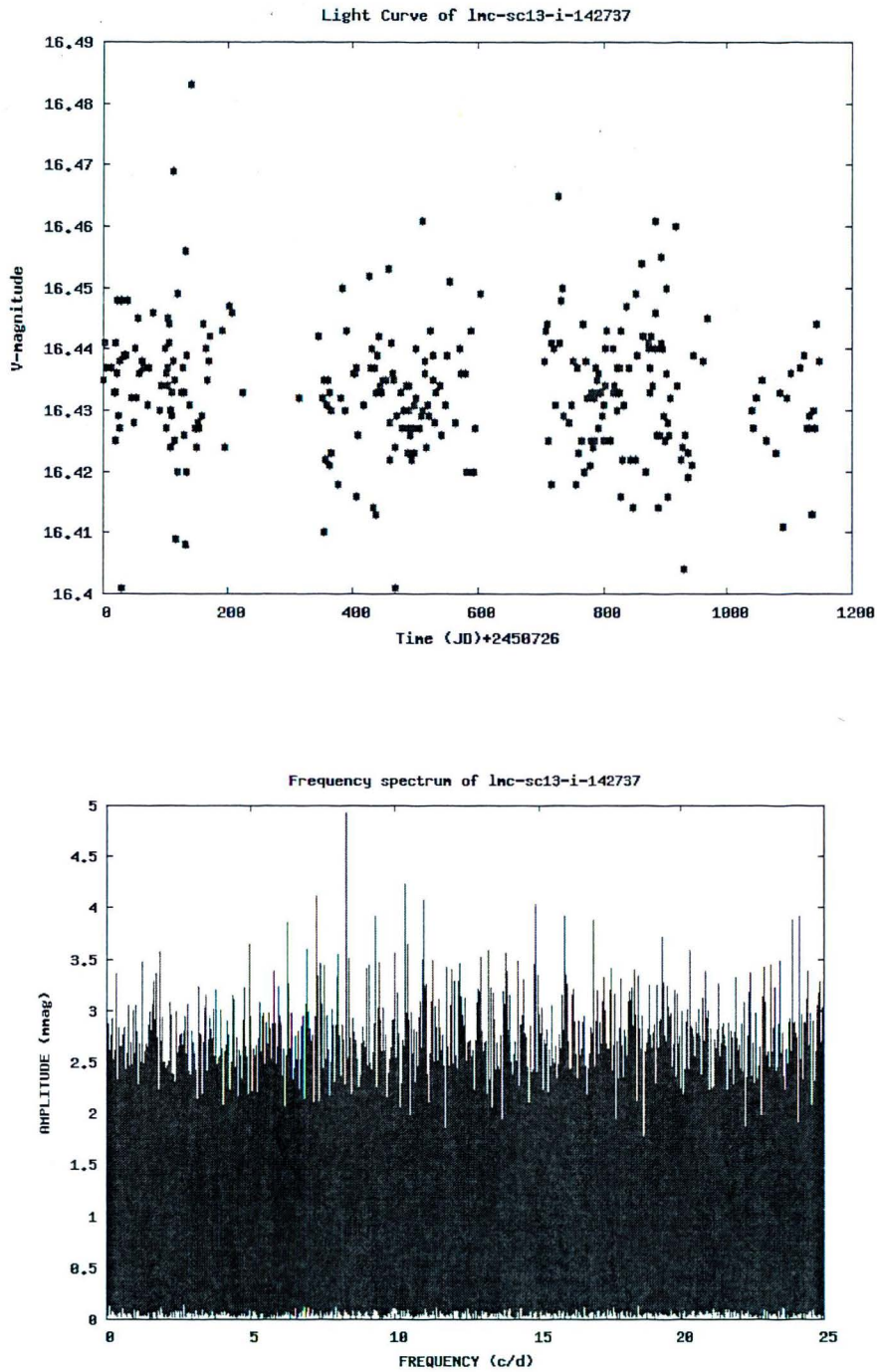


Figure 2.14: The top plot shows the OGLE light curve of lmc_sc13_i_142737, and the bottom plot is its frequency spectrum. The highest peak between 5 and 10 c/d looks promising. The other peaks look more like aliases and are as high as the noise level.

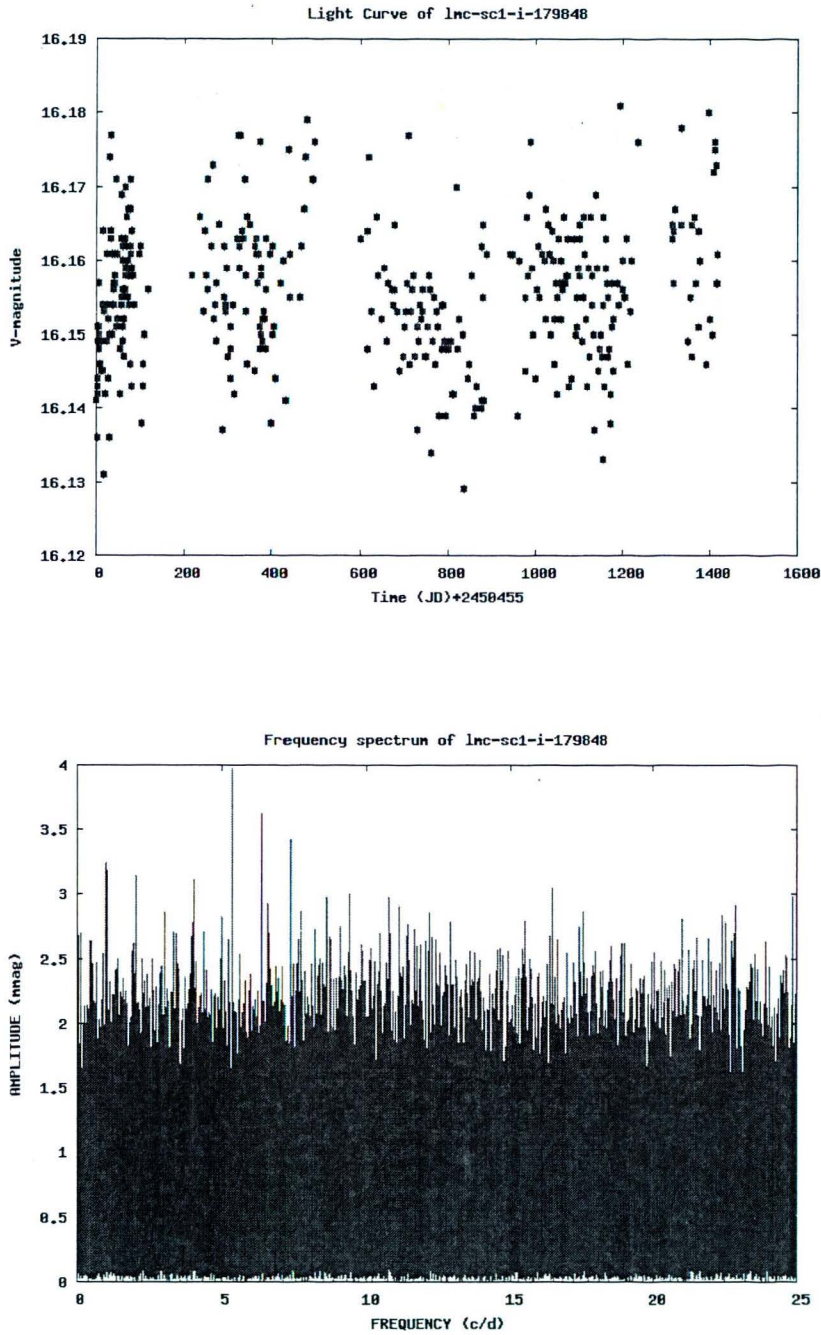


Figure 2.15: The top plot shows the OGLE light curve of lmc_sc1-i-179848, and the bottom plot is its frequency spectrum with a Nyquist frequency of 6.369 c/d. The highest peak at around 5 c/d looks promising. The other peaks on the right of the highest peak look like aliases given the noticeable gaps on the light curve.

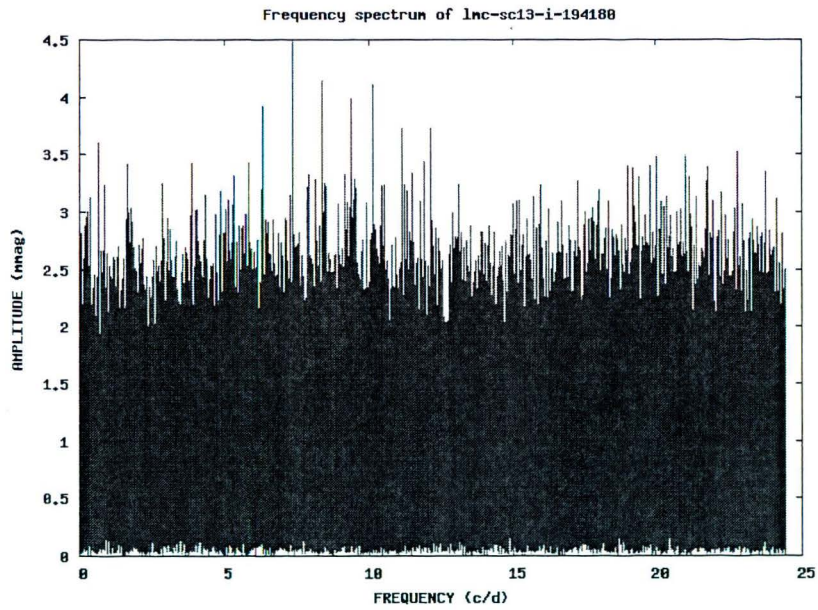
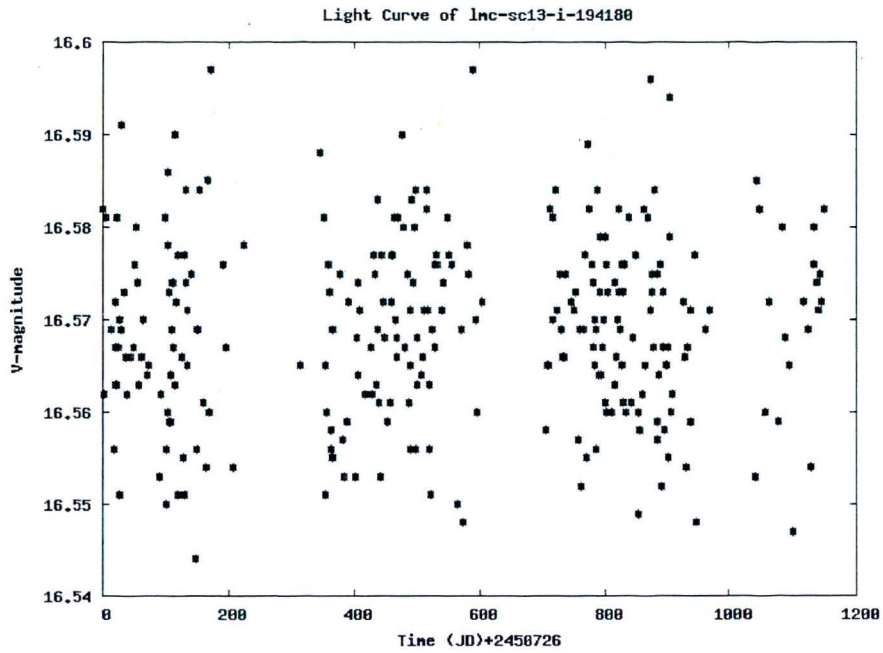


Figure 2.16: The top plot shows the OGLE light curve of lmc_sc13_i_194180, and the bottom plot is its frequency spectrum. The highest peak between 5 and 10 c/d looks promising. The other highest peaks might be aliases and the other noticable feature is the noise level of about 3.3 mmag.

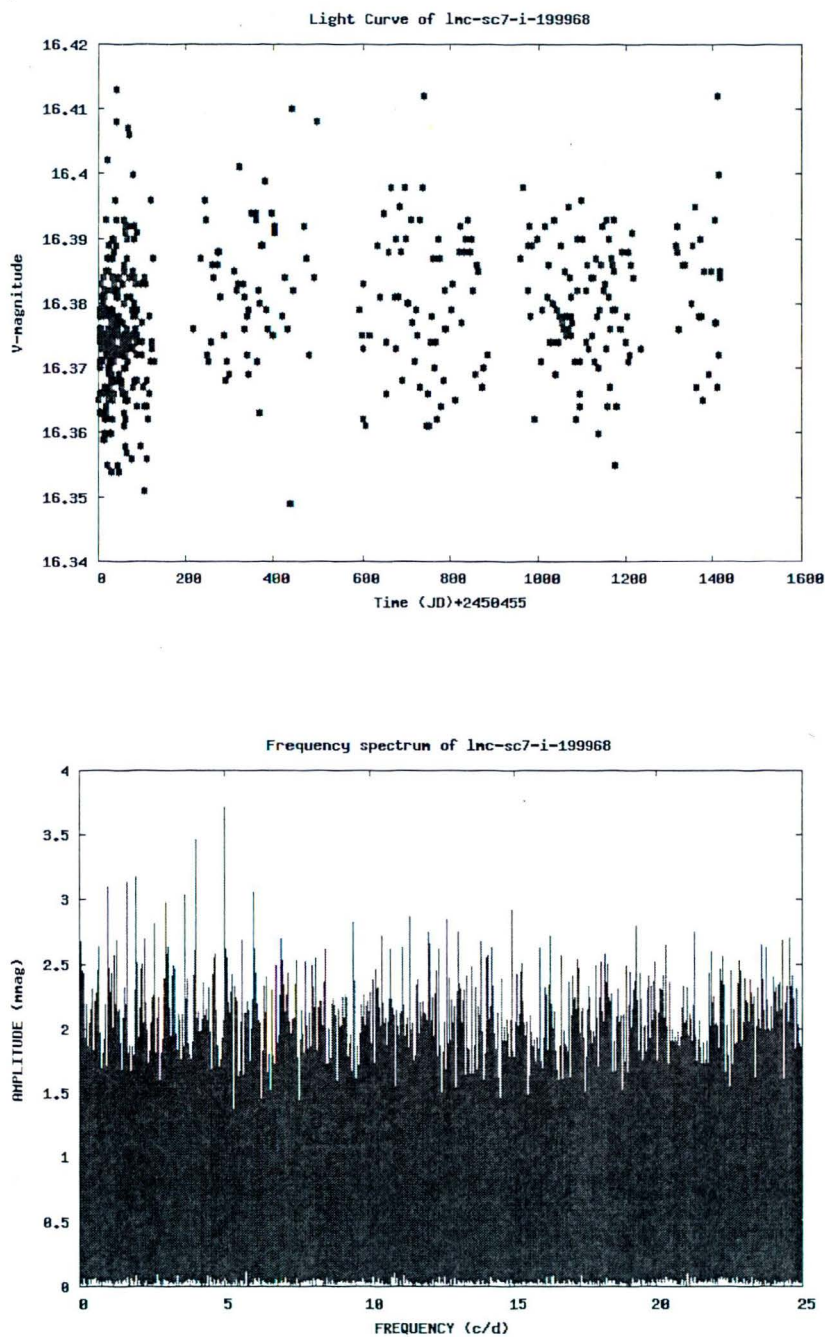


Figure 2.17: The top plot shows the OGLE light curve of lmc_sc7_i_199968, and the bottom plot is its frequency spectrum with a Nyquist frequency of 9.338 c/d. The highest peak between 0 and 5 c/d looks promising. The other peaks look like aliases. Also notice the noise level of about 2.8 mmag on the spectrum.

The amount of observing time allocated to us was very limited, we therefore had to reduce our target list further to six. This was done based on the amplitude of the highest peaks on the power spectra. The highest amplitude candidates from the 16 candidates mentioned above were given more priority because that was going to make it easy to detect such frequencies when doing the reductions from the observed data. From this final selection, six targets were given a priority for November 2010, February 2011 and March 2011 observation campaigns. Table 2.2 shows the properties of six selected targets. Due to their long OGLE II names shown on column 1 of Table 2.2, these targets were given shorter names also shown on column 2 of Table 2.2.

Star Name (LMC-)	shortened name	$\alpha(2000.0)$	$\delta(2000.0)$	V (mag)	V-I (mag)	B-V (mag)
SC10-i-212836	T11	05 11 32.5	-68 43 30.5	16.371	-0.062	-0.102
SC10-i-132705	T4	05 11 0.8	-68 50 30.3	16.334	-0.023	-0.088
SC13-i-194180	T9	05 06 31.5	-68 25 20.7	16.392	-0.170	-0.160
SC13-i-142737	T6	05 06 27.1	-69 06 17.4	16.423	-0.003	-0.048
SC13-i-218743	T13	05 07 5.0	-68 58 20.6	16.123	-0.012	-0.076
SC3-i-108137	T1	05 27 50.6	-69 26 25.4	16.468	-0.100	-0.199

Table 2.2: A table showing properties of six candidates selected for the November 2010, February 2011 and March 2011 observation campaigns. These properties are positions in right ascension and declination, and their photometric indices.

The OGLE maps of the targets on Table 2.2 are shown in Fig. 2.18. The targets are pointed with an arrow on each map.

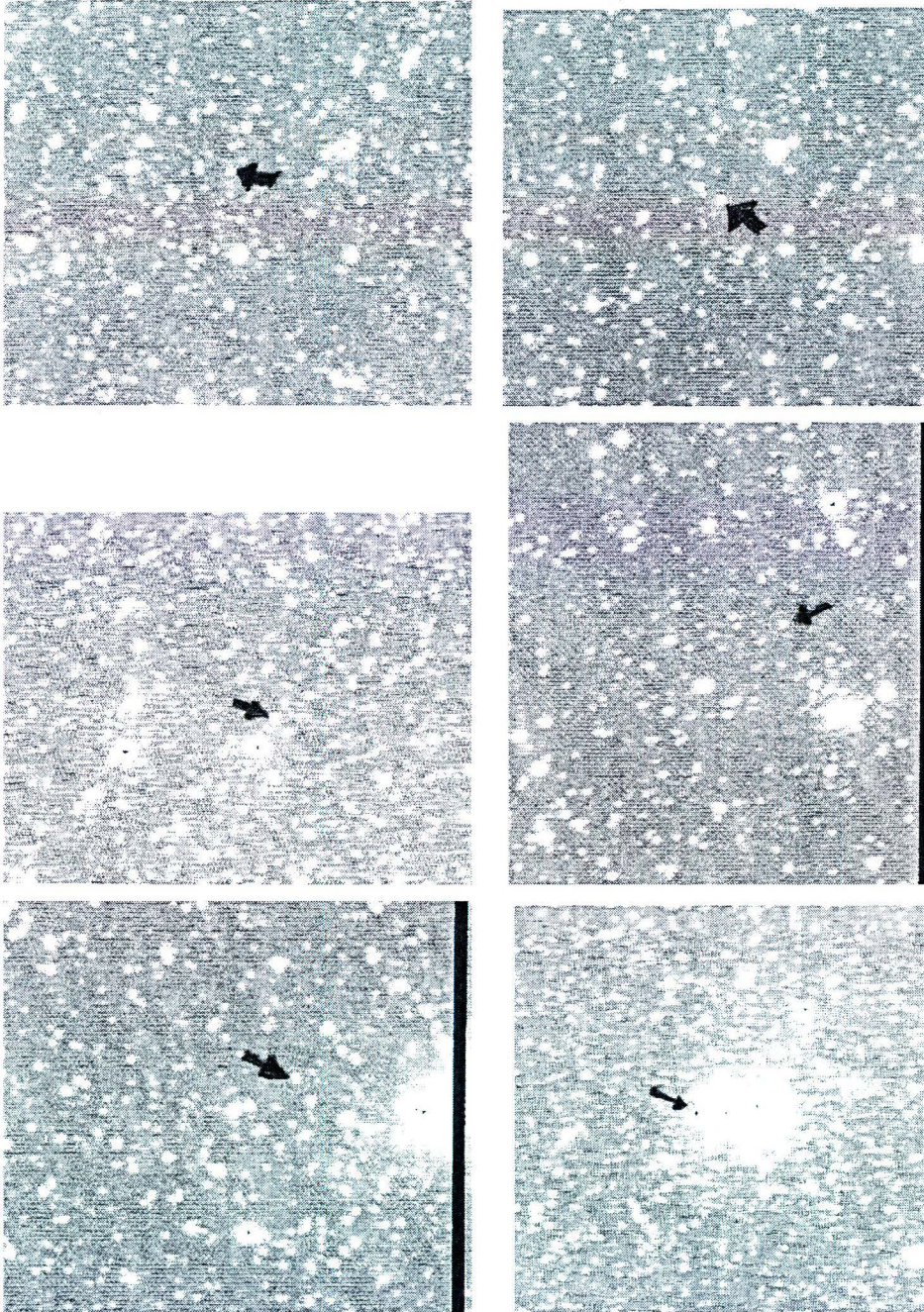


Figure 2.18: A figure showing the OGLE finding charts of all six β Cephei candidates selected for observations, from top left, clockwise is the image of T1, T4, T9, T13, T11, and T6. Each candidate is indicated by an arrow on the images (see <http://ogledb.astrouw.edu.pl/ogle/photdb/>).

Since the final six β Cephei candidates were selected based on their FAP, frequency of maximum peak and the amplitude corresponding to the frequency of maximum peak, the values of these quantities are shown on Table 2.3. It can be seen from column 4 of this table that all six candidates have FAP value less than 0.05 and the frequencies of their maximum amplitudes in column 2 are within the pulsation frequency range of β Cephei stars.

Star name	Frequency (c/d)	Amplitude (mmag)	FAP
T11	6.2497	4.1075	0.0393
T4	8.0253	4.7778	0.0050
T9	7.3806	4.4839	0.0332
T6	8.3250	4.9208	0.0072
T13	14.8984	9.9077	0.0192
T1	4.8689	13.9289	0.0000

Table 2.3: A table showing the frequency, amplitude and FAP of the six β Cephei candidates selected for observations.

2.2 Observations

The targets were observed in three campaigns spanning five months, the following part of the chapter deals with each campaign. The section 2.2.1 gives details of the November 2010 campaign followed by the February 2011 campaign in section 2.2.3 and finally the March 2011 campaign in section 2.2.4.

2.2.1 The November 2010 campaign

The first observation of two targets, LMC-SC3-i-108137 (T1) and LMC-SC13-i-218743 (T13), was conducted from the 3rd until 24th of November 2010. The targets were observed using *UBVI* Johnson's filters on the 1.0-m SAAO telescope (a brief description of this telescope is given in section A.5 of the Appendix), which was equipped with the STE4 CCD detector. The observations ran for three weeks. The following paragraph explains the CCD detector used during our observation runs.

The STE4 is a CCD camera (also known as the SAAO CCD instrument) used as a detector in either 1.0-m or 1.9-m telescopes, it has 1024×1024 pixels and is normally enclosed within a cryostat which is mounted at the cassegrain focus of the telescope. When this detector is mounted on the 1.0-m telescope, as it was during our observations, it is controlled by a Linux-based computer and this might also be the case for when it is used on 1.9-m telescope. Other properties of the STE4 detector are given in Table 2.4.

Manufacturers	SITE
Chip type	Back-illuminated
Number of pixels	1024×1024
Pixel size	24 microns square
Scale (1.0-m)	0.31 arcsec/pix
Field of view (1.0-m)	317×317 arcsec squared
Readout noise	6.5 electrons
Readout time	43 sec.
Scale factor	2.8 electrons/ADU

Table 2.4: A table showing the properties of the STE4 CCD, it can be seen that some of the properties are specific to the 1.0-m telescope (those written 1.0-m in brackets) (see <http://www.sao.ac.za/facilities/instruments/sao-ccd>).

During the first week both targets were observed every night. T13 was observed first, after being observed through four filters mentioned above, the telescope was pointed to T1 to also observe it through those four filters. We noticed after one week that we were gathering very few data points per star per night, and thus the second and third weeks were used to observe one star per night in order to get enough data points per star. The number of data points depends on the length of the integration time per filter, and as it can be seen on the observation program in Table 2.5 below, one observation cycle per target took 21 minutes plus the readout time, which depends on the binning in use. The binning concept is explained on section 2.2.2 below. For (1×1) binning, the readout time is a bit longer than (2×2) binning. Table 2.4 shows the readout time for (1×1) binning, which drops to about 15 seconds for (2×2) binning.

The integration times shown in Table 2.5 were estimated by trial and error in such a way that all filters gave enough counts (> 20000 ADU). We also tried to get equal counts from all filters even though we did not succeed because STE4 CCD detector has different sensitivity on each four filters we used. The CCD is more sensitive in I filter

Filter	Exposure time (sec)
U	480
B	240
V	240
I	300

Table 2.5: A table showing integration time used for each Johnson's filter employed during November 2010 observation run.

and least sensitive on U filter, therefore our attempt to get equal counts from all filter could have resulted in even longer integration time, resulting in few data points per night.

Given the variability period and amplitude of β Cephei stars, few data points were not enough to reveal significant visual variability on the light curve. Out of twenty-one nights we had for observations, only ten were photometric and other nights were either cloudy or humid. Table 2.6 shows the weather condition details during November 2010 observation nights.

Date (Nov,2010)	Target observed	Duration of Observation (hrs)	Atmospheric conditions
05	T13	2.07	wind speed = 16 km/h humidity = 69 %
06	T13 & T1	3.5	wind speed = 40 km/h humidity = 70 %
07	T13 & T1	5.9	wind speed = 10 km/h humidity = 50 %
09	T13 & T1	0.9	wind speed = 50 km/h humidity = 90 %
10	T13	6.2	-
12	T1	5.8	-
13	T13	3.8	wind speed < 20 km/h Thin clouds at horizons
14	T1	5.25	Bad seeing wind speed > 50 km/h
17	T13	6.08	humidity = 81% wind speed = 10 km/h
18	T1	6.0	wind speed = 40 km/h
19	T13	1.08	Thick clouds wind speed > 40 km/h
20	T1	4.0	humidity = 49 % wind speed > 40 km/h

Table 2.6: A table showing the starting weather conditions during observations of November 2010 targets. The targets were given the names (T1, T13) during the target selection, however their OGLE names are: T1 = LMC-SC3-i-108137 and T13 = LMC-SC13-i-218743 (see <http://wasp.astro.keele.ac.uk/live/>)

2.2.2 Observing Procedure

On clear nights, it was necessary to take the flat field frames and this was strictly done at around ten minutes before the Sun sets. The times for sun set for each day were looked up from the Sutherland almanac. At this time the sky gave enough counts (≥ 10000 ADU) on the frame so that it can be used for flat fielding. The flat field frames are used for flat-fielding and trimming off the overscan region from the image. During certain nights flat field frames were taken in the morning when the sky was clear and neither too dark nor too bright. This interval was traced by exposing firstly in I-filter (or U-filter in the evening) and if the collected counts were below 60000 ADU but above 10000 ADU, then the sky was neither too bright nor dark. This also depended on how low towards the horizon one points and how long exposure was made. For our case, we pointed the telescope 30° above East horizon (for morning flat fields) or 30° above West horizon (for evening flat fields). In the evenings, the flat field frames were taken for all four Johnson's filters: UBVI, in the sequence shown on Table 2.7.

filter	int. time (sec.)	binning
U	0.5	(2×2)
B	0.5	(2×2)
V	0.5	(2×2)
I	2.0	(2×2)

Table 2.7: A table showing the filter sequence used during afternoon sessions for taking flat field frames, columns 2 and 3 show integration time per filter and binning, respectively.

The filter sequence on Table 2.7 was preferred because the sky gets darker (counts drop below 10000 ADU) in the U-filter first, then in the B and V filters and lastly in the I-filter. The morning sequence for taking flat field frames was the reverse of the sequence above, i.e (IVBU). Exposure times were selected so that much time was not being spent on one frame, it had to be as short as possible to avoid taking few frames during suitable flat field interval explained above. When the sky was still too bright in a certain filter for (2×2) binning, then (1×1) binning was deployed even though the

frames with (1×1) binning were not of importance for the observed target. This is because the targets were observed using (2×2) binning, so taking flat frames with (1×1) binning was just to trace the darkening/brightening of the sky so that (2×2) binning could be deployed.

The binning is explained as follows: If (1×1) binning is in use then the total ADU is determined by reading out the counts from one pixel at the time. If (2×2) binning is used, the total ADU is determined by adding counts from $2 \times 2 = 4$ pixels and read that out as one pixel output, e.g. if one pixel readout is 10 ADU then (2×2) binning gives 40 ADU and read that as output from one pixel. Therefore (1×1) binning during flat fielding was used to check when ADU was $\leq 15\,000$, so that (2×2) binning could be deployed to produce total ADU $\leq 60\,000$. Fig. 2.19 shows four flat field frames, one from each of UBVI filters, observed during the first night of observation.

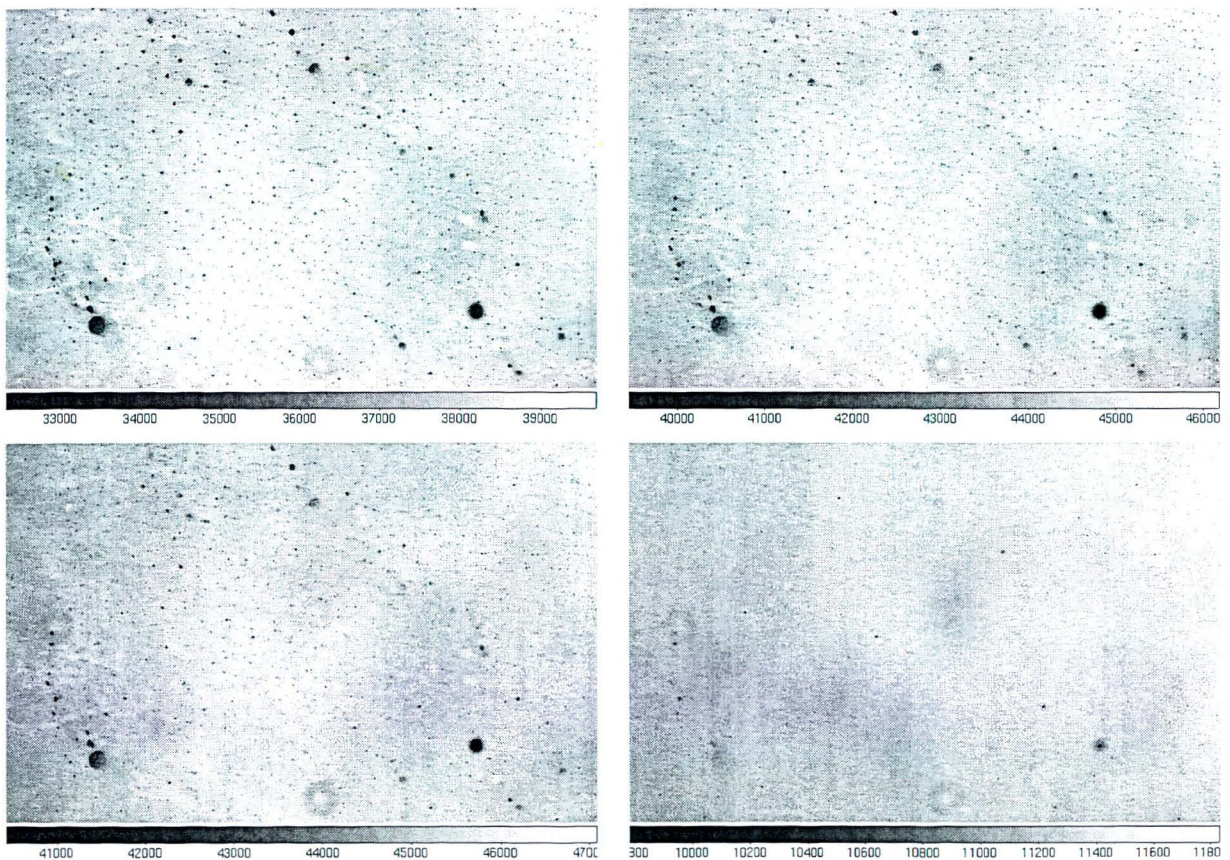


Figure 2.19: Figures showing the flat field frames from four Johnson filters (UBVI) obtained during the first observation run (05 November 2010). In clockwise direction, U-flat is the top-left frame, B-flat is top-right, I-flat is bottom-right and V-flat is bottom-left. The black spots on the frames are bad pixels which should be noted during observation, so that the target is not placed on any of those spots. The scale below each flat frame is the brightness scale and it increases from left to right.

During every observing session, it was very important to choose the guide star. This star was used to guide the telescope to point steadily to the same field for the rest of the observation session. The 1.0-m telescope has a CCD camera, called Acquisition & Guiding camera (from here on called the A/G camera), for monitoring the movement of this guide star and adjusting the telescope position accordingly, in order to point at the same field throughout the observation session. The guide stars for the two observed targets did not have to be the same star every night because any star outside the main CCD camera's field of view was suitable for guiding. Bright stars were used for guiding,

this gave an advantage when there were very thin clouds because bright stars can be observed through such clouds, whereas fainter stars cannot and that was going to cause the mis-guiding of the telescope, causing the telescope to drifts.

Since the fields we observed were too crowded (see Fig. 2.20), it was very important during each night to place the target star at the same place on the CCD frame. This is useful during data reduction since the target's position on the frame has to be specified, therefore having the target at various spots in the crowded field was going to confuse the data reduction program resulting in the wrong stars being mistaken with the target stars.

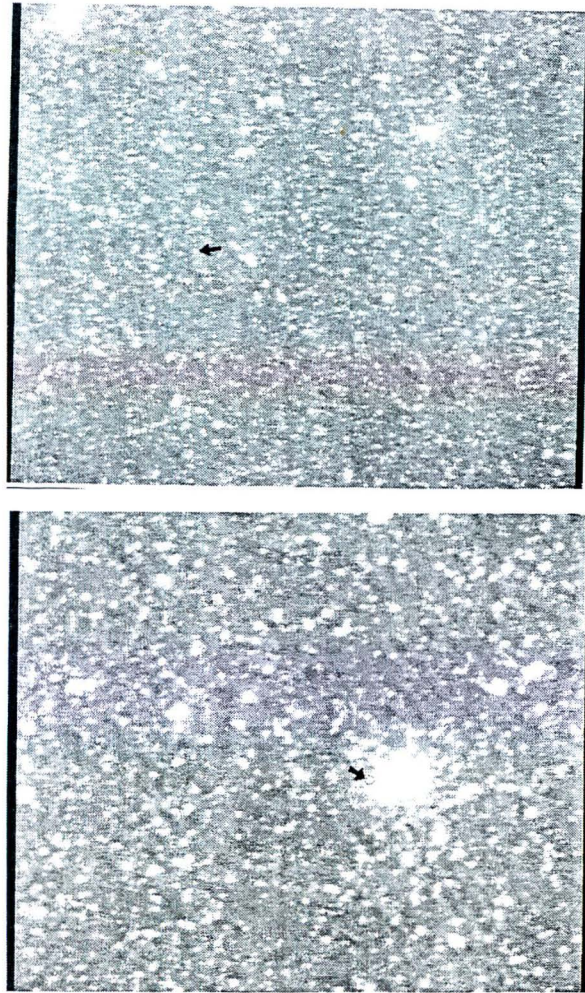


Figure 2.20: A figure showing the CCD frames of T1 (top) and T13 (bottom) from the LMC, observed during November 2010 observation run. T13 is located in the cluster (pointed by an arrow) seen on its frame and it can be seen that both frames are very crowded and T1 is the star pointed by an arrow towards the left of the frame.

Due to unavailability of finding charts for the targets observed, alternative finding charts had to be used and they are shown on Fig. 2.21. These were the finding charts of the stars close to our targets, this approach was convenient because the CCD instrument we used (STE4) has dimensions of 317×317 arcsecond squared field of view. The difference in position between finding chart stars and the targets are as shown on Table 2.8.

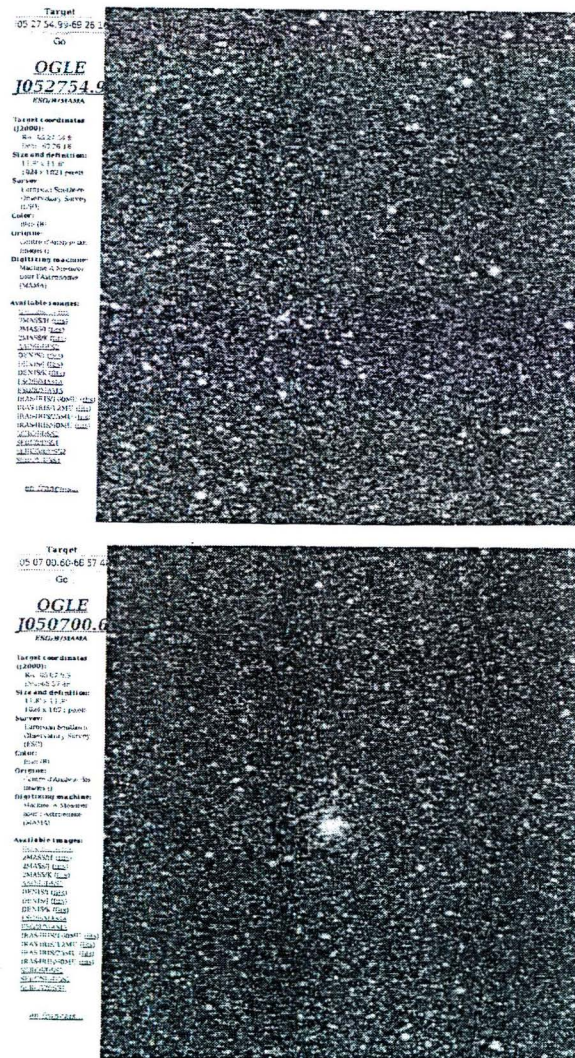


Figure 2.21: A figure showing the finding charts of the stars close to T1 (top) and T13 (bottom) that were used to locate our targets (see <http://simbak.cfa.harvard.edu/simbad/>).

Target star	Finding chart star	Difference in coordinates
T-1(LMC-SC3-108137)	LMC-SC3-108962	$\Delta\alpha = 00\ 00\ 03$
$\alpha(2000.0) = 05\ 27\ 52$	$\alpha(2000.0) = 05\ 27\ 55$	$\Delta\delta = 00\ 02\ 08$
$\delta(2000.0) = -69\ 28\ 08$	$\delta(2000.0) = -69\ 26\ 16$	
T-13(LMC-SC13-218743)	LMC-SC13-221205	$\Delta\alpha = 00\ 00\ 4.5$
$\alpha(2000.0) = 05\ 07\ 05$	$\alpha(2000.0) = 05\ 07\ 0.5$	$\Delta\delta = 00\ 01\ 27$
$\delta(2000.0) = -68\ 58\ 21$	$\delta(2000.0) = -68\ 57\ 48$	

Table 2.8: A table showing the coordinate differences between target stars and close-by stars whose finding charts were used to find the correct fields during observations of T13 and T1.

From the coordinate differences it can be seen that the targets are close to the finding chart stars and also within the STE4 frame. However, it was later found that there are maps on the OGLE web-page taken during OGLE II observation mission, these maps could be used as finding charts as well and as a matter of fact these maps were used during our third observation run.

The observation of targets began at around 22:30 South African Standard Time (SAST) even though the twilight was before 20:00 SAST and observations could run until the morning twilight, this was because the 1.0-m telescope can access LMC when it is at an altitude of about 35° above horizon during the first observation run. Although LMC is visible every night at Sutherland, it rises at different times every night due to the motion of the Earth around the Sun. The visibility curves of both T1 and T13 for the first observation run are shown in section A.3.1 of the Appendix.

2.2.3 The February 2011 campaign

During the second observation campaign, which started from the 9th until the 16th of February 2011, the target T13 was replaced with another target, LMC-SC13-i-142737 (T6) because T13 was found to be located within the cluster, which made it hard for reduction software to detect it for photometry. The two observed targets (T1 and T6) were getting high enough above the horizon to be accessed with 1.0-m telescope at around 20:30 SAST, and they were accessible until mid-night. This second observation session ran for seven nights, but due to bad weather conditions, only two nights were available for observations. Weather condition details for the entire session are shown on Table 2.9 below. The observing procedures were kept the same as in the first observation run, i.e. the same integration time in each filter and the same binning. This is because T6 is also a 16th magnitude star in V-filter and also is located on the same OGLE field (i.e.SC13) as T13.

Date (2011-02)	Target observed	Duration of Observation (hrs)	Atmospheric conditions
09	-	0	Heavy clouds and high humidity
10	-	0	Rain, heavy clouds and high humidity
11	T6	7	Thin clouds at East horizon
12	-	0	Thin clouds and high humidity
13	T1	6.0	High humidity at around 02:30 SAST
14	-	0	Heavy clouds
15	-	0	Heavy clouds and rain

Table 2.9: A table showing the weather conditions experienced during the February 2011 observation run, and the third column shows the number of hours spent on observing. It should be noted that by high humidity it means the humidity > 90 % and $T_{ext} - T_{Dew} < 1.5$ °C (see <http://wasp.astro.keele.ac.uk/live/>).

2.2.4 The March 2011 campaign

The third observation run started on the night of 9th March 2011 and ended two weeks later on the 23rd March 2011. This observation session was different from the other previous observations because, firstly only one target was observed for the entire run. We decided on this because T6 was not observed enough during the second run and also LMC was accessible only for about 2.5 hours. Secondly, T6 was observed in one filter only, the I-filter. The reason for using one filter was none other than the short time LMC could be observed. The idea was to get enough data points during that 2.5 hours so that we can detect the pulsation frequencies. The I-filter was preferred over other four (UBVR) because of high sensitivity of STE4 CCD on this filter, this was mentioned by Dr. Balona and was confirmed during data reductions where in most cases good signals of the targets were only found in I-filter, and in other filters the targets were too faint for reduction programs to pick them out. However, forcing the data reduction programs to select the targets is a matter of adjusting the thresholds as explained under the data reductions in chapter 3. We deployed 2×2 binning when observing T6 ($V = 16.423$) and exposed each frame for 300 seconds because T6 is faint. The weather conditions for the entire observation session are given in Table 2.10.

Date (2011-03)	Target observed	Duration of Observation (hrs)	Atmospheric conditions
09	T6	2.07	Wind speed = 10 km/h $T_{ext} - T_{Dew} = 19.8^{\circ}\text{C}$ Lightning and heavy clouds towards North-East horizon
10	T6	2.5	Wind speed = 20 km/h Thin clouds at horizon $T_{ext} - T_{Dew} = 20.9^{\circ}\text{C}$
11	T6	2.5	Wind speed = 23 km/h Very thin and scattered clouds layers $T_{ext} - T_{Dew} = 20^{\circ}\text{C}$
12	T6	2.5	Wind speed = 20 km/h Thin clouds towards west horizon $T_{ext} - T_{Dew} = 16.9^{\circ}\text{C}$
13	T6	2.5	Wind speed = 30 km/h $T_{ext} - T_{Dew} = 9.5^{\circ}\text{C}$ Thin clouds on East horizon
14	T6	2.5	Wind speed = 10 km/h $T_{ext} - T_{Dew} = 21.4^{\circ}\text{C}$ Thin clouds all over the sky
15	T6	2.5	Wind speed = 30 km/h $T_{ext} - T_{Dew} = 15.6^{\circ}\text{C}$

Table 2.10: A table showing weather conditions for the third observation run during observation of T6 (LMC-SC13-i-142737) whose coordinates are $\alpha(2000.0) = 05\ 06\ 27.1$ and $\delta(2000.0) = -69\ 06\ 17.4$. Column 4 shows the weather during the beginning of the night (see <http://wasp.astro.keele.ac.uk/live/>).

T6 was already high enough to be observed during the evening twilight, and was getting too low at around 22:30 SAST. The visibility curves of T6 are shown in section A.3.1 of the Appendix. Before the observations could be started, the zero points had to be reset before the telescope could be pointed at the targets. The setting of the zero points was to enhance the pointing of the telescope. The zero points had to be set for every field we had the target on. The zero points setting procedure is given in section A.5.1 of the Appendix. Fig. 2.22 shows the frame of T6 where this star's position on the frame is pointed by an arrow.



Figure 2.22: A figure showing the frame taken for T6, which is located towards the bottom-center of the frame and indicated by an arrow.

The coordinates and photometric indices of the stars we observed in the three campaigns are shown in Table 2.11.

Star name	RA $\alpha(2000.0)$	DEC $\delta(2000.0)$	V (mag.)	B (mag.)	V-I
T1	05 27 50.61	-69 26 25.4	16.468	16.268	-0.100
T6	05 06 27.10	-69 06 17.4	16.423	16.374	-0.003
T13	05 07 04.99	-68 58 20.6	16.123	16.047	-0.012

Table 2.11: A table showing the photometric properties of T1, T6 and T13. These were taken from the OGLE II database (see <http://ogledb.astrouw.edu.pl/>).

Chapter 3

Data Reductions

This chapter describes the reduction steps that were applied to the data from our three observation campaigns. As the fields are crowded, a precaution was required. Preliminary data calibration was performed at the telescope using the DUPHOT program written by Dr D. O'Donoghue of SAAO.

3.1 DUPHOT

Preliminary calibration of frames include bias subtraction and flat fielding. No separate bias frames are required, the bias calibration is obtained from the overscan region of the CCD. The first step is to run `scanfits` which simply reads the FITS headers and classifies each frame as a target or flat field frame. A `scanfits` produces two files containing names of flat field frames (`flt.tmp`) and target frames (`obj.tmp`). The filter number and binning factor associated with each frame are also stored in these files since each frame observed with a given filter and binning factor needs to be divided by the flat field for the same filter and binning factor. The next step is to correct each frame for the bias and apply flat fields. This is accomplished by running `Cleen`. This program reads the names of the flat field frames from `flt.tmp`, removes the bias from each frame using the overscan region and combines flat fields for each filter and binning factor, and creates a normalized flat field for that filter and binning factor. The files thus created have names of the form `flatp01.fg`, where

p: The binning factor. If the binning is 1×1 , then $p = 1$ and if the binning is 2×2 then $p = 2$, etc.

f: The filter number in wheel A.

g: The filter number in wheel B.

In our case, the following normalized flat fields were produced: `flat201.18`, `flat201.28`, `flat201.38` and `flat201.58` for the U, B, V and I filters respectively. These normalized flat fields must be present before the target frames can be calibrated. The flat fields produced in this way are an average of the individual flat fields collected during a particular observing run. Since the condition of the filters, telescope and optics changes with time, it is not advisable to combine flat fields from different runs. Once the normalized flat fields have been created, they can be applied to the target frames. `Cleen` reads the target frame from the file `obj.tmp`, removes the bias and divides by the corresponding flat field frame. It also uses interpolation to fix bad pixels. The calibrated target frames have names `cnnnnmmmm.fits`, where:

c: identifies the file as a calibrated frame;

nnn: the run number;

mmmm: the frame number.

The programs, `scanfits` and `Cleen`, complete the calibration process. Preliminary reductions are possible at the telescope, these involve careful inspection of the frames and creation of a file called `windows`, which contains the positions of the targets for which photometry is required. Firstly, the images are viewed using `ds9`. Fig. 3.1 shows an example of such an image opened with `ds9`.

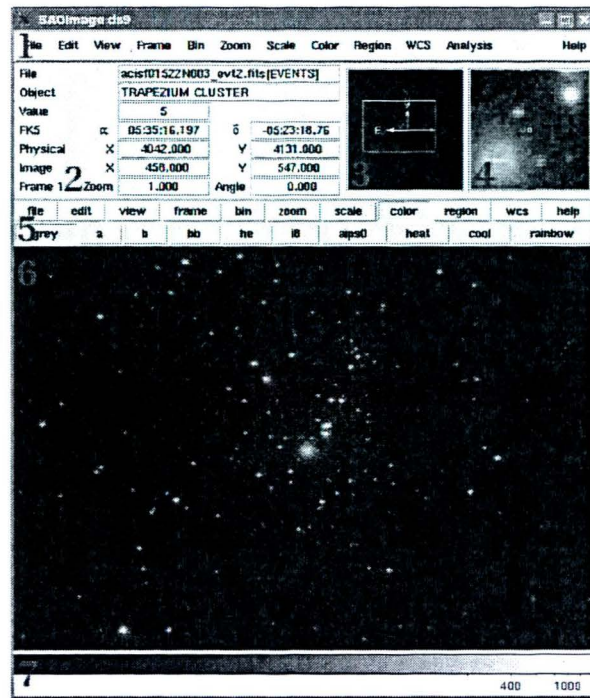


Figure 3.1: A figure showing the ds9 interface and how it displays the FITS images (see: <http://hea-www.harvard.edu/RD/ds9/user/gui/index.html#components>).

To obtain photometry at the telescope, the positions of each target are found by placing the cursor on the target and writing the displayed coordinates to the `windows` file. This file consists of numbers in the format

```
x1 y1 b1
x2 y2 b2
. . .
. . .
xi yi bi
```

where

xi: the column number of the target;

yi: the row number of the target;

bi: the full-width half maximum (FWHM) of the Point Spread Function (PSF) to be fitted (PSF fitting is described on section 3.2.1).

It should be noted that the calibrated frames need to be used when creating the `windows` file and not the raw frames. Once the `windows` file has been created, one can run `Reduce`. This program needs several inputs as described below.

- The directory containing the calibrated target frames.
- The run number, i.e. `nnn` in `cnnnnnnnn.fits`.
- The number of the first frame to be reduced. i.e. `mmmm` in `cnnnnnnnn.fits`.
- The name of the CCD that was used. There is a list of available names and one selects the number associated with the name. In our case the CCD name is `STE4`.
- Specify whether the data being reduced was collected using a single filter or more than one filter. For a single filter one types 0 and for more than one filter one types 1.
- The filter number and the corresponding normalized flat field name.
- The thresholds, and there are two kinds of thresholds required. Those are:

The threshold above the sky brightness below which no star will be measured, and it is represented by T_{min} ; and

The threshold above the sky brightness for fitting the PSF to stars, this is represented by C_{min} .

- Choose between the windowed mode and non-windowed mode. The option here depends on whether one chose to create a `windows` file as explained above or not. For the windowed mode, the `windows` file should be present in the working directory.
- Select the crowded field mode.
- The aperture size.

After this information has been entered, `Reduce` will run `DoPhot` (see section 3.2) as a background job and produce files of the form `sum.nnnnnnnnn`, which we will call

"sum" files. These files contain the positions, profile-fitted magnitudes and aperture magnitudes of each star.

Next, the program `Mergesum` is executed. This program extracts the data for the same star from all the sum files and produces a file called `datin.dat` which contains the times, profile-fitted and aperture magnitudes for each star. At this point, the data reduction is complete and one can plot the light curve for each star and even calculate the periodograms. However, the magnitudes produced at this point are raw magnitudes in which each frame has the same zero point. Because frames were obtained at different air masses and different sky conditions, the zero point varies from frame to frame. To eliminate this variation we use the difference in magnitude between the target star and one or more comparison stars rather than the raw magnitude. This technique is called *differential photometry*.

To apply differential photometry, a set of suitable comparison stars needs to be selected for each field. A simple way of doing this is to plot unstandardized magnitudes for each star and select those stars which have the same variation with time. Once these stars have been selected, one runs `diffphot`. This program simply finds the mean magnitude of the comparison stars and subtracts it from the target star. It requires, as input, the numbers of the comparison stars that have been selected and produces a file called `answer.dat` in the same format as `datin.dat`. Fig. 3.2 shows two plots, one produced from `datin.dat` and showing the raw magnitudes and the other produced from `answer.dat` showing differential photometry. Even though the above descriptions of reduction programs sound straight forward and efficient, it was not easy to apply them on our frames. This was because of the crowded fields and faintness (> 16 mag.) of our targets. These two factors posed a severe challenge to these programs. At first we thought that the high minimum threshold (T_{min}) value might have been the reason for our lack of success, but even after pushing it down to as low as 50 ADU, our targets were only detected in I-frames. It was therefore necessary to tailor `DoPhot` to our needs and the following section gives its description.

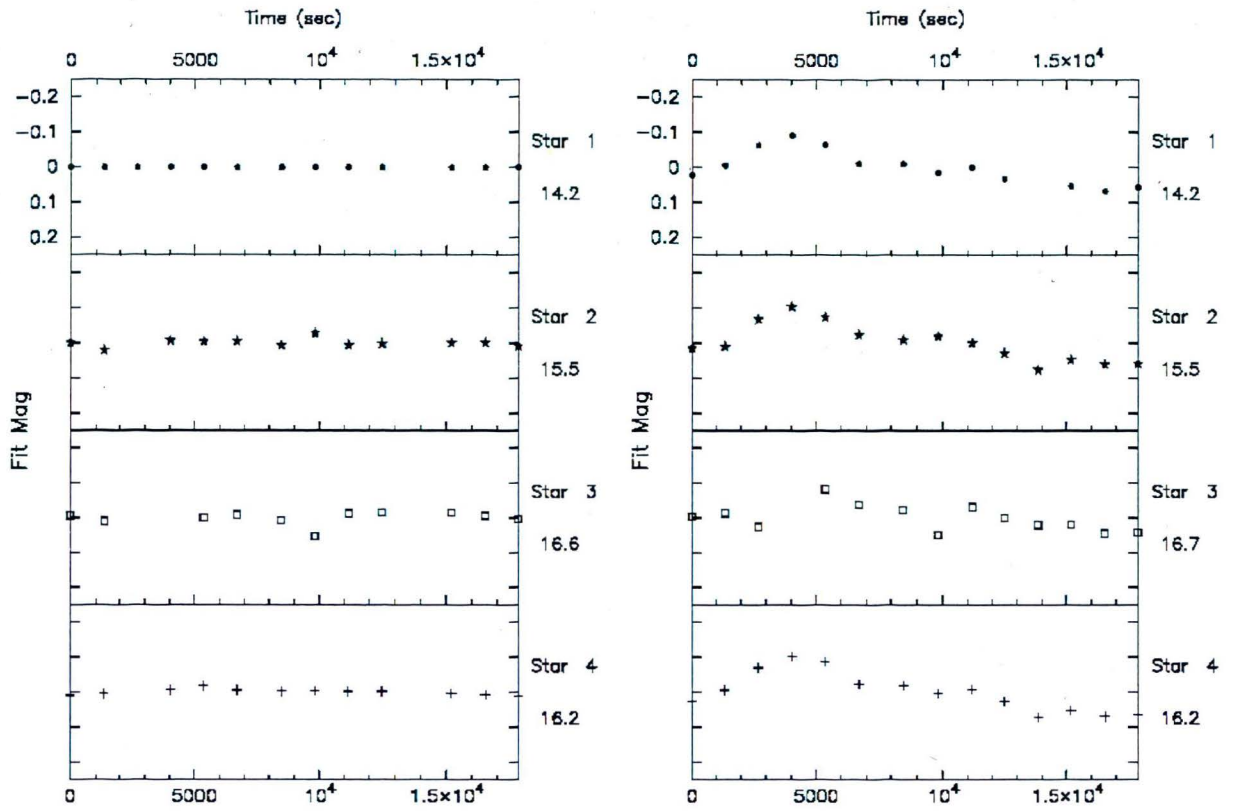


Figure 3.2: A figure showing the light curves of first four brightest stars from the results of November 2010 observation campaign. On the right panel are the raw light curves and on the left panel are light curves after differential photometry has been applied, and star 1 was used as a comparison to demonstrate an effect of differential photometry on light curves of other three stars as targets.

3.2 DoPhot

DoPhot (Schechter et al. 1993) is a program for automatically identifying stars on a CCD frame and obtaining profile-fitted magnitudes as well as aperture magnitudes. The output from DoPhot also includes the type of object (star, galaxy, close double, cosmic ray, saturated pixel, etc.), its position in pixel coordinates, and fitted parameters. In this section we describe PSF fitting and aperture photometry as performed by DoPhot. To facilitate the use of DoPhot, several ancillary programs are used and these are also described.

3.2.1 Point Spread Function (PSF) fitting Photometry

The PSF is the distribution of the light from a point source over a two dimensional surface, in our case the distribution of photons over a CCD detector. There are a few factors why light from a point source is distributed on the CCD detector. Atmospheric turbulence scatters photons from the point source. The limited resolution and aberrations in the optics of the telescope and camera further spread out the light. Finally the limited resolution of the CCD detector and imperfections cause further deterioration of the image. The left panel of Fig. 3.3 shows the PSF above the atmosphere and the right panel is the PSF on the ground. The broadening of the PSF from the left panel to the right panel is caused by the above mentioned factors.

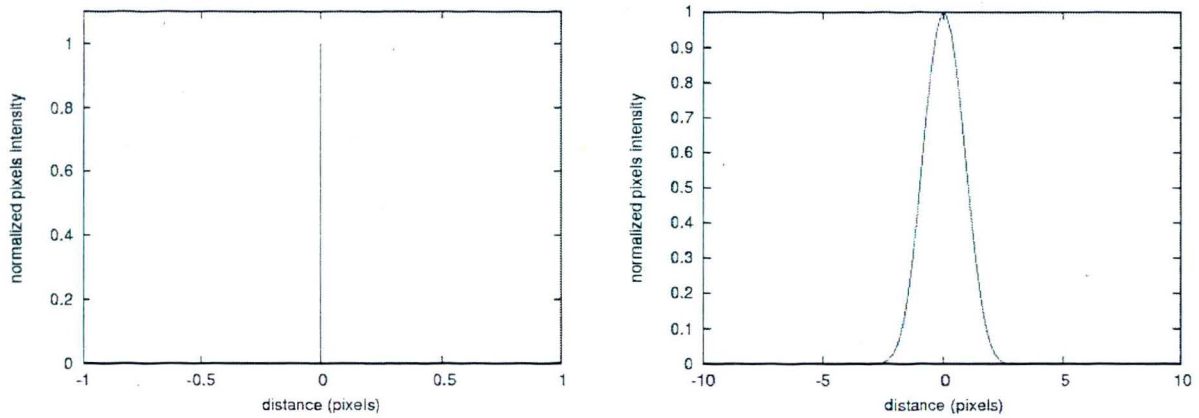


Figure 3.3: Plots of intensity as a function of position in CCD (in pixels) from above the Earth's atmosphere (left panel) and from the Earth's surface (right panel). These plots represent atmospheric and instrumental effects on the light from a point source star.

In aperture photometry, which is described in section 3.2.2 below, one is always faced with the problem of how to select the optimal aperture size and sky annulus size. Since the sky background is given equal weight as the star image when summing the total intensity in an aperture, the resulting measure of the intensity is not optimal. To resolve this problem, it is necessary to use PSF fitting. In this technique a given profile $P(x, y)$ is fitted to the image of the star and surrounding sky background and the brightness of the star is determined by one of the adjustable parameters. To use this method, an accurate analytical or empirical function to represent $P(x, y)$ is required. This is found by using a few of the brightest stars on the image.

In DoPhot it is assumed that the PSF remains the same for all objects on the CCD except for the sky background and a scaling factor, which determines the height of the PSF, is proportional to the brightness of the star. The PSF for a star is not easily described by any analytical function and will vary from frame to frame. For ease of calculations, however, it is desirable to represent the PSF by a suitable function containing a sufficient number of free parameters which can be adjusted to obtain a good fit to the PSF. In DoPhot, the following expression is used:

$$P(x, y) = A_1 + \frac{A_4}{1 + t + \frac{1}{2}t^2 + \frac{1}{12}t^3} \quad (3.1)$$

where

$$t = \frac{1}{2} \left(\frac{X^2}{A_5} + 2A_6XY + \frac{Y^2}{A_7} \right) \quad (3.2)$$

and

$$X = (x - A_2) \quad (3.3)$$

$$Y = (y - A_3) \quad (3.4)$$

where x and y are pixel positions and A_1 , A_2 , A_3 , A_4 , A_5 , A_6 and A_7 are adjustable parameters.

To determine the optimal shape parameters, A_5 , A_6 , A_7 , DoPhot only uses the brightest stars on a frame. The threshold for a star to be included in the determination of the shape parameters is one of the parameters required as input to DoPhot. Stars used in this way are given a type classification of 1. For stars with peak intensities below this threshold, the shape parameters are fixed and the only parameters that are adjusted are A_1 , A_2 , A_3 and A_4 . These fainter stars are classified as type 7.

In DoPhot one requires an initial estimate of the shape parameters A_5 , A_6 , A_7 as well as the threshold level described above. This is achieved by setting initial threshold level quite high. On the first pass only stars exceeding this level are detected and fitted. The initial estimates of the shape parameters are refined iteratively for best fit. Once a good fit has been obtained, the stars identified above this threshold are removed from the image in such a way that they can be restored exactly to the initial state at a later stage. Integer arithmetic is used for this purpose. When stars are removed, they are replaced with corresponding noise to preserve the statistics. With the stars above the initial threshold removed, the threshold is lowered and further stars identified and fitted. These stars are again removed and the threshold lowered. The process continues until the threshold has been reached for shape parameter fitting. At this point DoPhot

has a list of the positions and approximate parameters for all stars above the shape-fitting threshold. DoPhot then replaces the stars that it has removed and performs a full simultaneous fit to obtain the best values for A_5 , A_6 , A_7 . DoPhot then re-computes the remaining parameters for each star using the optimal shape parameters as known values. As mentioned above, the stars used in this procedure are given a classification of type 1.

Once type 1 stars are fitted and removed, the process is continued below the shape-parameter threshold in the same way. This time, however, the shape parameters are held at their optimal values and only A_1 , A_2 , A_3 and A_4 are determined. Stars below the shape-parameter threshold are classified as type 7. One of the input parameters is the minimum threshold above sky background. Once this threshold has been reached, DoPhot stops.

There are several problems which DoPhot needs to consider. One of them is that there may be two or more stars very close to each other. DoPhot attempts to fit close doubles by a combination of two PSFs and, if it succeeds, the resulting object is given a classification type 3. If no fit can be found, a classification type 4 is assigned. If the object is too broad to be fitted by the analytical PSF, it is assumed to be a galaxy (type 2). Other types are also recognized. A full list of types is given in Table 3.1.

3.2.2 Aperture Photometry

Aperture photometry simply measures the light within a given aperture which will include the combined intensity from star and sky. For CCD aperture photometry, it is convenient to use a square centered on a specific pixel coordinate and of given width. The measurements of light within the aperture is done by adding up the intensities of all pixels inside the chosen aperture.

The main consideration when choosing the size of the aperture is to avoid as far as possible light from the surrounding sky but, at the same time, sample as much light from the star of interest as possible. This means that too large an aperture should be

Classification type	Object
1	PSF star (bright)
2	galaxy
3	close doubles
4	unconverged star
5	too few active pixels
6	decommissioned object
7	faint star
8	saturated pixels

Table 3.1: A table showing DoPhot classification types and the objects for each type.

avoided, not only because it includes too much sky, but also because there is a high possibility of including light from neighbouring stars. On the other hand, too small an aperture might result in omitting a significant amount of light from the star. A good choice of suitable aperture size is to choose an aperture which just encloses the visible edge in a faint star.

Since light in the aperture includes the sky as well as star, it is necessary to determine the contribution from the sky. For this purpose, an annulus is defined around the aperture and the average intensity per pixel for the sky is calculated. A reasonable size for the annulus is about 4 or 5 times the size of the the inner aperture. The magnitude of the star inside the aperture is obtained by taking the logarithm of the difference between total intensity within the inner aperture and intensity of the sky background. This is mathematically described below.

If I_{ij} is the intensity on pixel (i, j) , then the total intensity in the inner aperture is:

$$I_T = \sum I_{ij} \quad (3.5)$$

and if B is an approximated sky background intensity, which is defined as:

$$B = \frac{1}{N_A} \sum b_{ij} \quad (3.6)$$

where N_A and b_{ij} are number of pixels inside the outer aperture (but outside the inner aperture) and the intensity on each pixels respectively. Then intensity of the star's brightness is given by

$$I = I_T - B \quad (3.7)$$

and therefore the magnitude of the star is given by

$$m_I = -2.5 \log_{10}(I) \quad (3.8)$$

In DoPhot, one needs to specify the aperture size. The sky annulus is automatically determined. During the course of the night, the seeing will change and one should adjust the aperture size to compensate for this and DoPhot does aperture size adjustment automatically. As already mentioned, differential photometry measures the difference in brightness relative to the mean brightness of a number of selected comparison stars. This removes changes in zero point caused by extinction and transparency variations. The profile fitting magnitudes are better suited to this type of photometry, as already mentioned.

On the other hand, aperture photometry must be used when it is required to measure the magnitude of stars obtained in different frames. For example, if one wants to obtain the true *UBVRI* magnitudes and colours of stars in a particular field, it is necessary to use aperture photometry of the field and use *the same aperture size* to measure E-region stars on another frame. It is necessary to use aperture photometry because the PSF shape parameters are different in different frames and therefore the height of the PSF is no longer proportional to the brightness of the stars. Therefore only aperture photometry can be used for this purpose. It is necessary to keep the aperture size the same so that exactly the same proportion of star light is measured in field stars and in the E-region stars.

For this reason, it is important to consider the best aperture size to be used for a certain number of frames and to keep the aperture size the same for those frames. Thus it is not always a good idea to allow the aperture size to vary according to the size of the image. It is best to choose an aperture which is too large rather than too small such that close to 100 percent of starlight is sampled by the aperture. The aperture size in this case is tailored to ensuring that the same proportion of light is sampled.

3.2.3 Preliminary steps

The frames to be reduced by `DoPhot` must be calibrated frames, i.e. the bias must be removed and the flat field applied. As already mentioned, the programs `scanfits` and `Cleen` were used to perform these functions. An example of the frame before and after calibration is shown in Fig. 3.4. The darker regions on the left and right edges of the image in the left panel are the overscan regions used to determine the bias level. These are removed, as can be seen in the right panel.

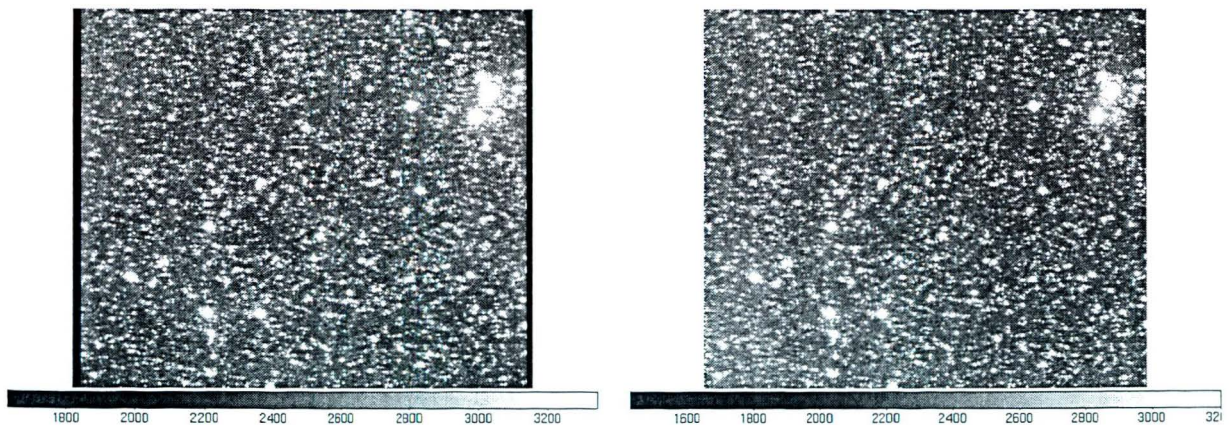


Figure 3.4: A figure showing two CCD images of the LMC-T6 field, the image on left panel is the raw image and the right panel image is flat-fielded and bias removed. Bias regions are the black regions on left and right edges of the raw image.

`DoPhot` requires information about the CCD and such information is stored in a file called `dophot.inp`. This file contains adjustable parameters specific to an instrument used for observations. The contents of `dophot.inp` are shown in section A.4.1 of the

Appendix. The first step in the reduction process is to create a list of file names for the FITS files to be analyzed. The program `ccdlist` reads the list from a file and asks for the aperture size to be used. The information is stored in a file `ccd.lis` which simply contains the file name (if the aperture size is to be determined by DoPhot itself) or the file name and aperture size (if the aperture size is fixed for all frames).

Next, the program `ccddophot` is run. This program takes the necessary information from `dophot.inp` and `ccd.lis` files. It writes such information to file `ccd.bat` which can be read by DoPhot. One of the options is how the sky should be modeled. There are three choices: PLANE, POWER and HUBBLE. The PLANE option means that the sky brightness is assumed to be uniform, however it can change linearly across the CCD frame in the form: $\text{sky} = a_0 + a_1 x + a_2 y$, where x and y are the pixel coordinates. POWER simply means that the sky brightness obeys a power law. The HUBBLE sky algorithm is mostly used for the photometry of stars in globular clusters (Schechter et al. 1993). In our case we always used the PLANE law. After this step, DoPhot can be run by typing:

```
dophot ccd.bat
```

As already mentioned, results are written to files with names such as `sum.nnnmmmm`. Here `nnn` and `mmmm` have the same meaning as explained earlier under `duphot` reductions. The contents of these files are described in section A.4.2 of the Appendix.

3.2.4 Extracting the photometry

The star numbering in the `sum` files is arbitrary so that the same number in different files is likely to refer to different stars. It is necessary to match the positions between different files and to re-number the stars so that the same number refers to the same star. This task is accomplished by the program `ccdcatlog` which takes a `sum` file as a template and calculates the transformation that is required to match each `sum` file to the template. This is accomplished by pattern matching using similar triangles. `ccdcatlog`

uses a minimum of five stars to determine the transformation. The output is the same as the `sum` files, but the stars are now correctly numbered in each frame. The output file names are of the form `cat.nnnnnnnnn` with the same format as the corresponding `sum.nnnnnnnnn` file. Only stars with types 1 and 7 are extracted.

Next, the data for each star in the `cat` files needs to be extracted so that the times and magnitudes are together in the same file. This is accomplished by running a program `ccdpick`. This program extracts the heliocentric Julian day from the header of the `cat` files. It converts the intensities given by the fitting parameter A_4 to magnitude, and also converts the aperture intensity to magnitude. Description of the output file is given in section A.4.3 of the Appendix.

At this point the reduction process is complete. However, the magnitudes are not differential magnitudes and suffer from airmass and transparency variations. These zero-point differences may be eliminated by selecting a suitable number of comparison stars which one assumes are not intrinsically variable. Choosing these non-variable stars can be facilitated by checking whether they are flagged as variables by another reduction program called `ISIS` (to be described later). It is also possible to select comparison stars by trial and error, and inspect selected stars visually. For each frame, one calculates the mean magnitude of all the comparison stars and subtracts this mean from all the stars. The program `ccdstd` performs this function. The program asks for the numbers of stars to be used as comparisons. The output file is `ccdstd.out` and has exactly the same format as `ccdpick.out` except that the columns for both profile fitting and an aperture magnitudes are now the magnitude differences.

3.3 ISIS

ISIS is a computer program for reducing CCD images using the image subtraction method. Alard & Lupton (1999) described this approach as optimal for dealing with images of different seeings. The image subtraction method requires an accurate convolution kernel which can be derived from a simple least square analysis of all the pixels of subtracted images. ISIS reduces images in three essential steps, and these steps are (i) image registration, (ii) image subtraction (which is the main reduction step) and (iii) photometric reductions. Each of these steps can be performed individually or together in a single step by executing a script called `process.csh` on the command line.

One needs to create a file containing the heliocentric Julian date (HJD) of each frame and an estimate of the typical image size in that frame. This information is stored in the `dates` file. One also needs to select a few frames with best image quality that will be used to construct a reference image. The names of these files are stored in file `ref_lis`. The ISIS scripts need to be edited prior to execution. For example, one needs to specify the location where the calibrated FITS files are to be found. This is done by editing the `process_config` file. Other configuration files (`default_config`, `phot_config`) may also need to be edited.

In order to understand the details of these three reduction steps, we will describe the individual steps.

3.3.1 Image Registration

In this step the images are corrected for astrometric errors such as shifts and possible small rotations. These corrections are made by selecting a reference image. This image is taken as a template which will be used for calculating transformations between itself and all other images. In DoPhot these transformations are made after the photometry has been extracted, but in ISIS this is an essential first step.

Image registration involves spline interpolation of the raw images and is executed by running the script `interp.csh`. This program produces registered and interpolated images corresponding to the raw images. The resulting files are prefixed by the name `interp`. The file `log_interp2` contains the residuals after astrometric transformation in both the X and Y directions and the number of stars used to calculate astrometric corrections for each image.

Next, one needs to create the reference image (this not the reference image that is chosen on the basis of its best seeing). This image is created by combining selected good seeing images into one reference image. The image combination is done in such a way that bad pixels, such as cosmic rays which will render the reference image useless, are rejected to avoid a build-up of defects. The program to execute this function is called `ref.csh`. The resulting reference image is called `ref.fits`. After the reference image is created, the process of image registration is complete.

3.3.2 Image Subtraction

Image subtraction is done by convoluting the reference image, `ref.fits`, so as to produce the closest possible match with the image being reduced. The convolved image and the image being reduced are then subtracted. Stars which are constant (and therefore have the same brightness on the two images) disappear. The operation of image subtraction method is controlled by the configuration file `default_config` which may need to be edited. However, we did not change anything in this configuration file. This image subtraction method is done by running the script `subtract.csh`. The mean standard deviation and the scatter values for each image is stored in file `log_subtract`. Also created are images with prefixes `kt` and `kc`.

During image subtraction, all non-variables on the images disappear and only the variables remain. Before photometry of these remaining variables can be done, they should be detected and found. The script which detects the variables from subtracted images is called `detect.csh`. This script produces two images called `var.fits` and

`abs.fits`. The image `var.fits` enables one to judge the level of variability of each star, the more intense the image, the higher its variability. A threshold value of the variability can be estimated from the intensity value of the images. This threshold will be used to extract photometry and produce light curves of the variables that lie above this threshold. The threshold value estimated in this way must be written to `process_config` under the SIG THRESH parameter.

The remaining step is to find all the variable stars with a variability greater than the specified threshold. Extracting these variables is done by running the script `find.csh` which writes the positions of the detected variables and their intensities in file `phot.data`.

3.3.3 Photometry

This last step of the ISIS reduction procedure produces the light curves for the variables in `phot.data`. Here the flux of the variables is determined by fitting the PSF to the images (PSF fitting photometry details are already discussed in section 3.2.1 above), therefore it is important that one looks inside the configuration file, `phot_config`, and enters an appropriate value for the size of the PSF to be fitted. The other parameters which require to be specified are the radii of the circle within which the pixels of the objects will be fitted for flux calculation and for the flux normalization. The above three radii are represented by parameters: `rad1.bg` and `rad2.bg` for sky background estimation, `rad_phot` for flux calculation and `rad_aper` for flux normalization. The script which creates the light curves is `phot.csh`.

The light curves are identified with names similar to the number of the variable in `phot.data`. Their names are of the form `lci.data`, where `i` is the number of the variable. These light curves contain six columns where the first column is the heliocentric Julian date as given in the `dates` file. The second column is the weighted flux and the third column is the standard deviation. The fourth and fifth columns contain the same parameters as the second and third columns, respectively. The last column is the instantaneous intensity of the image.

The light curves can be viewed by using any plotting program. The next chapter gives the results and also shows the comparison of the output light curves from DoPhot and ISIS reductions.

Chapter 4

Analysis and Results

This chapter describes how the outputs from the reductions were analyzed. The results of analysis and their discussions are also given.

4.1 Comparison of DoPhot and ISIS outputs

The DUPHOT programs encountered severe challenges in producing the final results because the targets were faint and located in crowded fields, hence DoPhot was deployed. We used DoPhot and ISIS to reduce Johnson-I frames of November 2010 and March 2011 campaigns. The idea was to determine the capability of DoPhot to reduce the crowded field frames by comparing its outputs to that of ISIS, which is designed to deal with crowded field frames.

Even though both programs managed to reduce the frames successfully, it was not obvious how to compare their outputs because DoPhot returns the PSF fit and aperture photometry while ISIS is returning the weighted *flux*. The ISIS output is somehow not yet understandable to us because at first we thought that the returned values are fluxes. But that posed another challenge when trying to convert those *flux* values to magnitudes. The challenge was that some values were negative, meaning that equation (3.8) in the data reduction chapter was not applicable. However, we made the light curves of four stars from DoPhot output and also four light curves of exactly the same

stars from ISIS output. In order to be certain that we selected the same stars from both DoPhot and ISIS outputs, we wrote a python program (see section A.2 of the Appendix) to match stars using (x,y) frame coordinates.

The light curves of the first matched stars are shown on the figures below. These are the light curves of the LMC_T13 and LMC_T6 fields. The LMC_T6 field frames are presented here because this field was observed in one filter only. These light curves are shown on the Figs. 4.1, 4.2, 4.3, 4.4, 4.5, 4.6, 4.7, 4.8 for the November 2010 campaign and Figs. 4.9, 4.10, 4.11, 4.12, 4.13, 4.14, 4.15, 4.16 for the March 2011 campaign. Table 4.1 shows the stars from DoPhot reductions and their corresponding stars from ISIS reductions.

DoPhot star ID	ISIS star ID
8	1
1885	2
343	4
7	6
6844	274
6942	844
6324	1113
6848	1842

Table 4.1: A table showing the identities of stars from DoPhot reduction outputs and their corresponding identities from ISIS reduction outputs.

November 2010 light curves

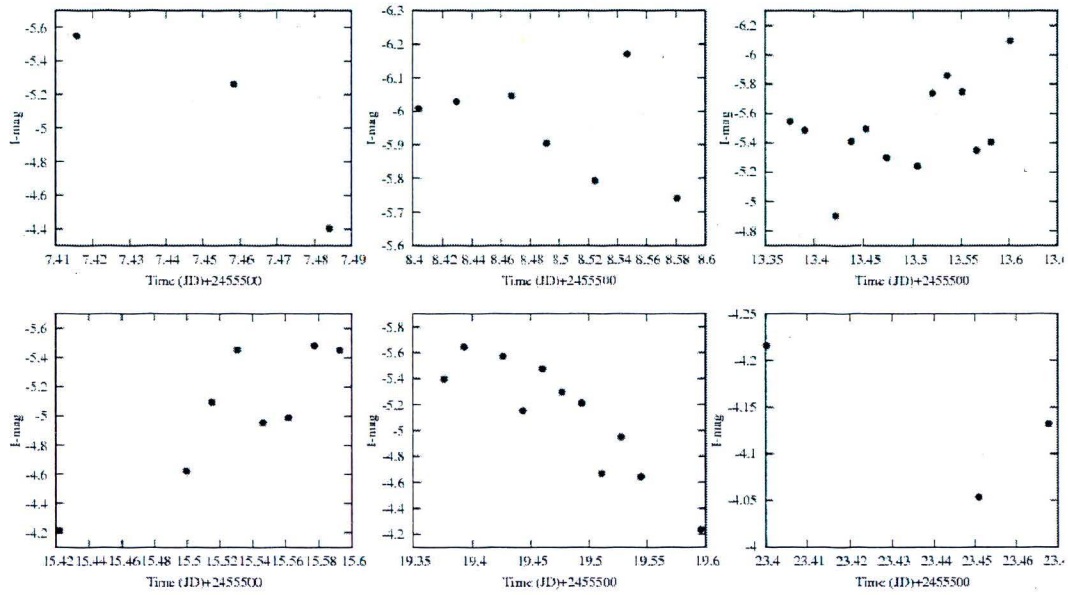


Figure 4.1: A figure showing the light curves of star 8 from DoPhot reduction outputs in different nights.

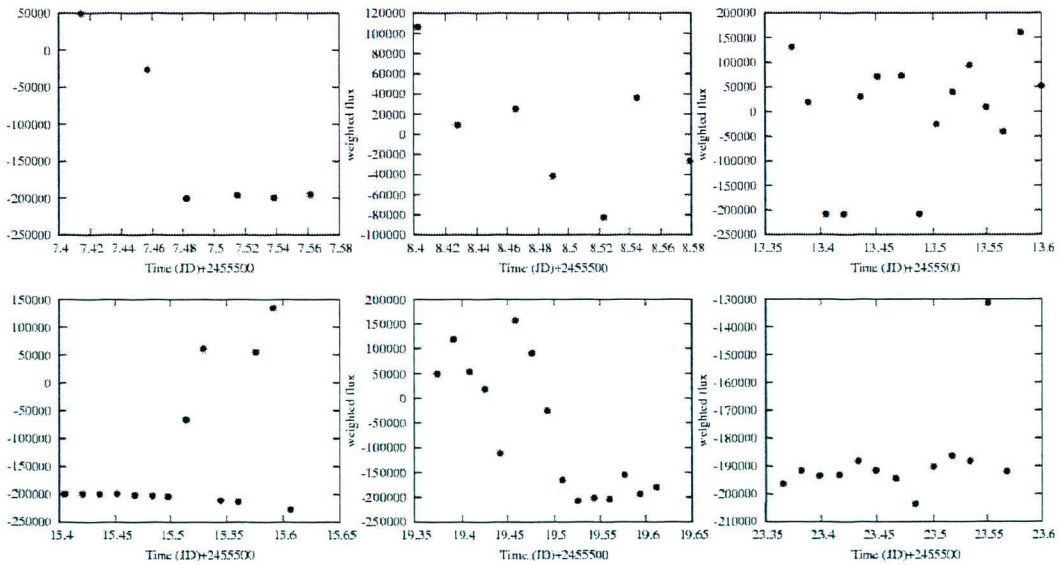


Figure 4.2: A figure showing the light curves of star 1 from ISIS reduction outputs, this is the same star as in Fig.4.1.

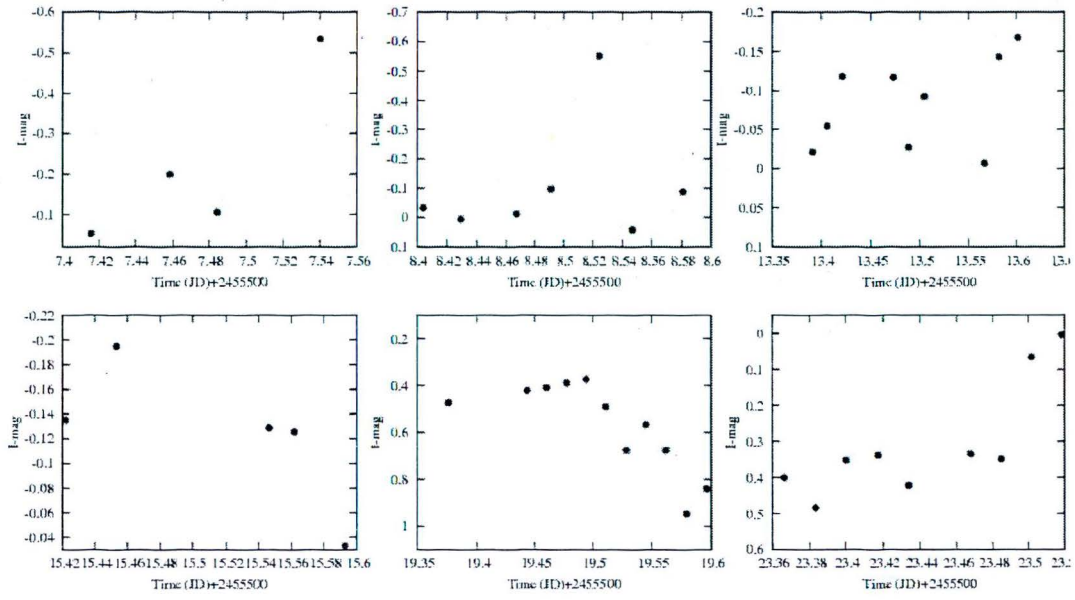


Figure 4.3: A figure showing the light curves of star 1885 from DoPhot reduction outputs in different nights.

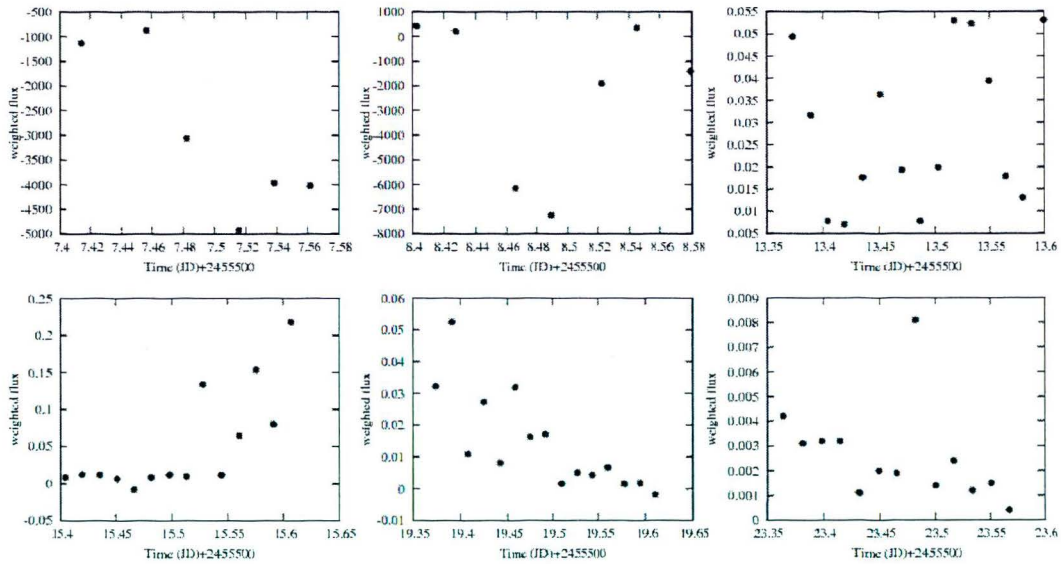


Figure 4.4: A figure showing the light curves of star 2 from ISIS reduction outputs, this is the same star as in Fig.4.3.

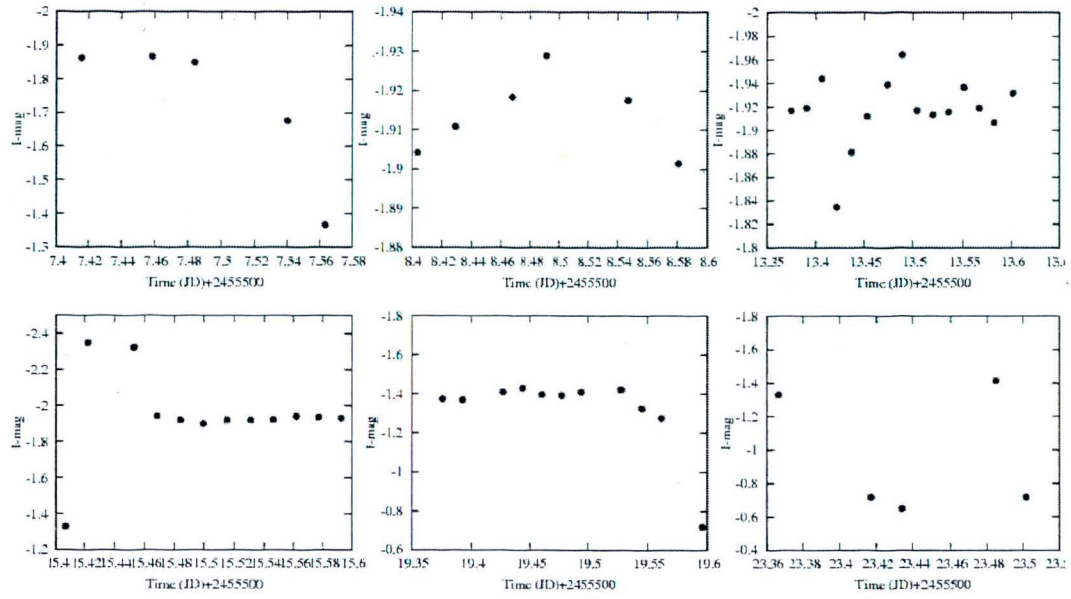


Figure 4.5: A figure showing the light curves of star 343 from DoPhot reduction outputs in different nights.

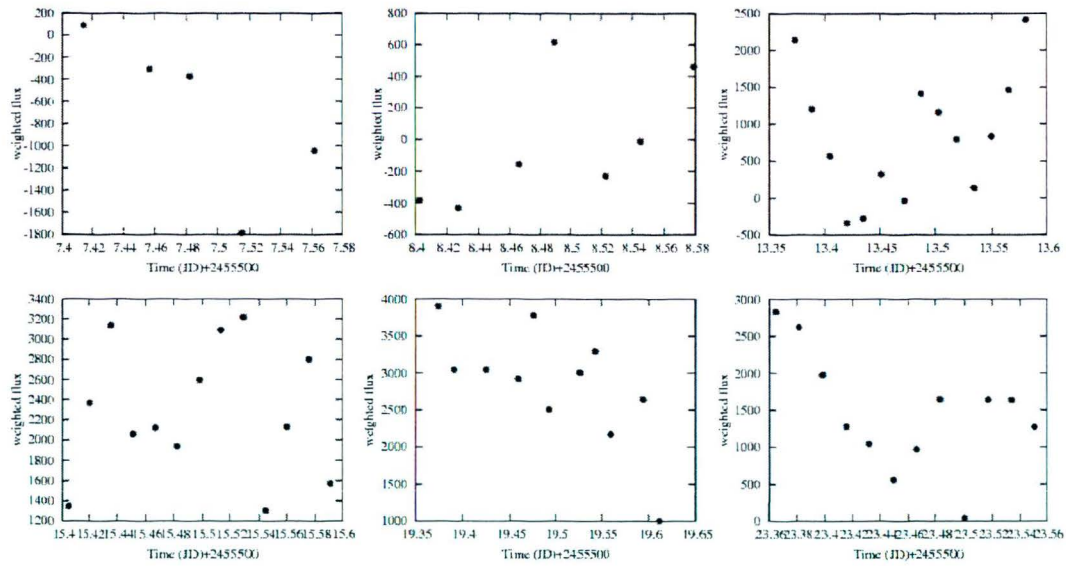


Figure 4.6: A figure showing the light curves of star 4 from ISIS reduction outputs, this is the same star as in Fig.4.5.

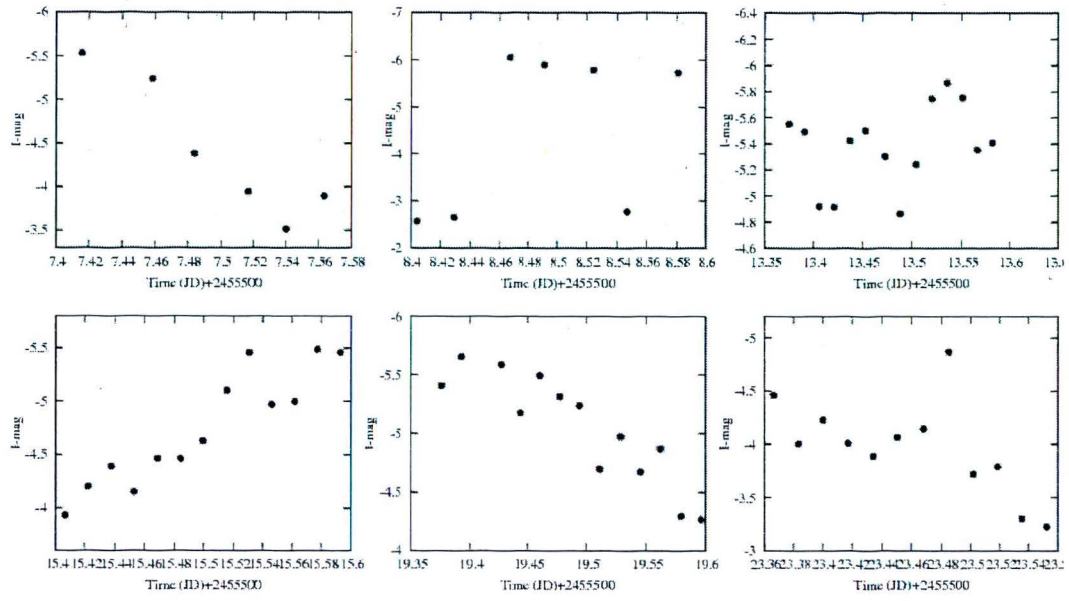


Figure 4.7: A figure showing the light curves of star 7 from DoPhot reduction outputs in different nights.

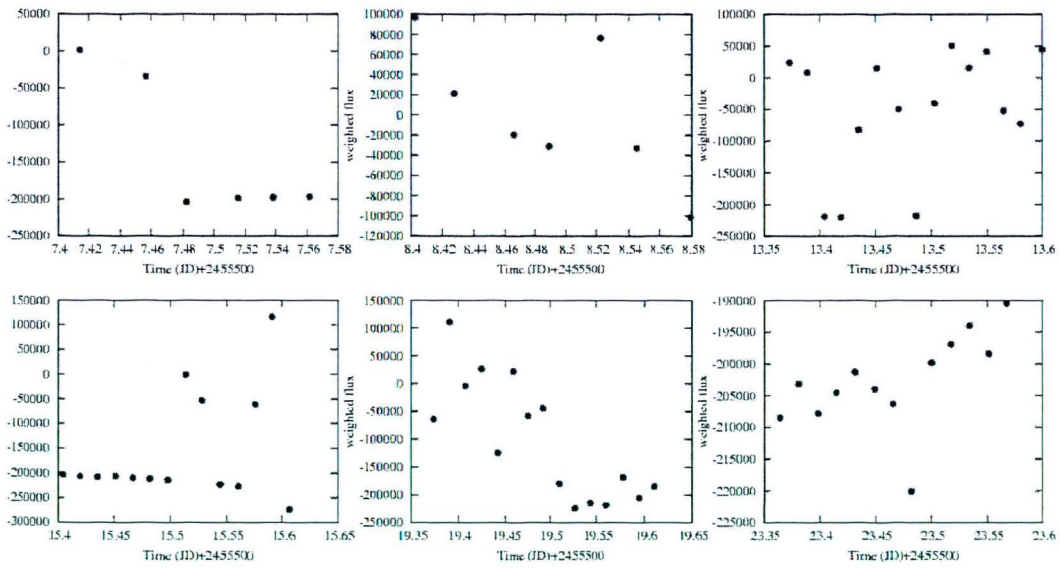


Figure 4.8: A figure showing the light curves of star 6 from ISIS reduction outputs, this is the same star as in Fig.4.7.

March 2011 light curves

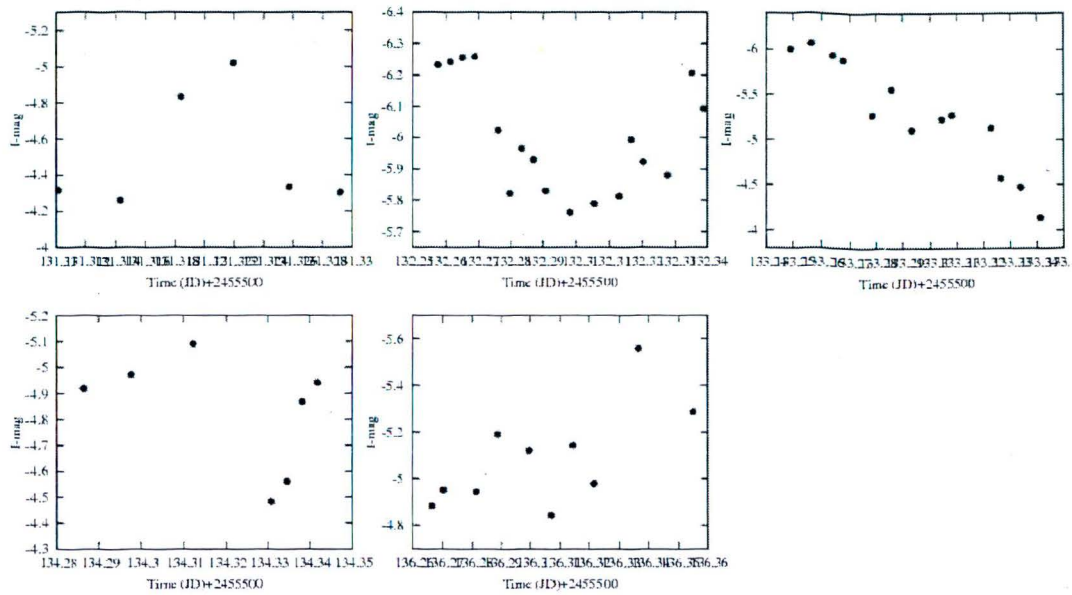


Figure 4.9: A figure showing the light curves of star 6844 from DoPhot reduction outputs in different nights.

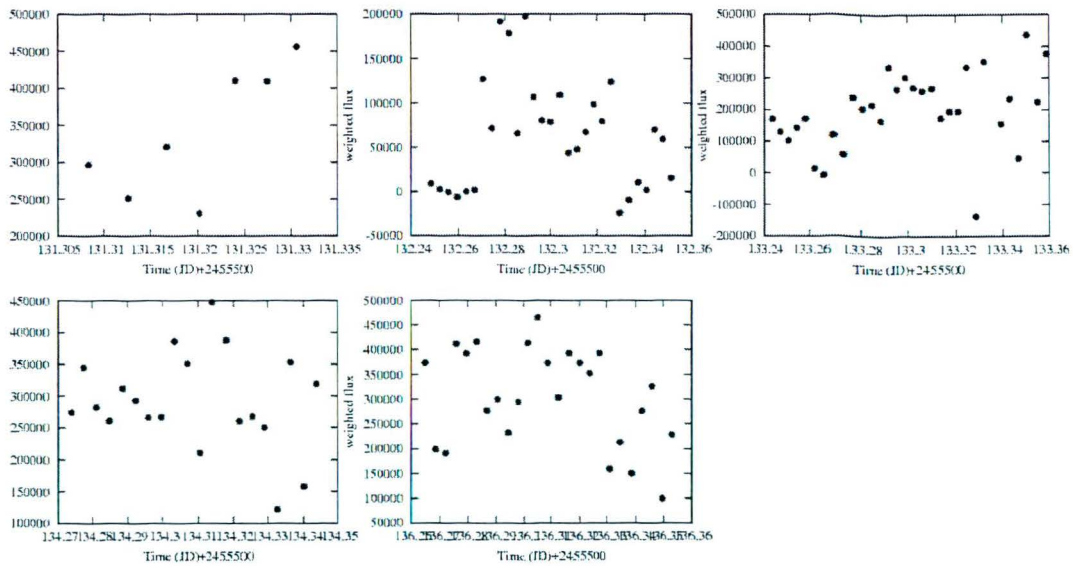


Figure 4.10: A figure showing the light curves of star 274 from ISIS reduction outputs, this is the same star as in Fig.4.9.

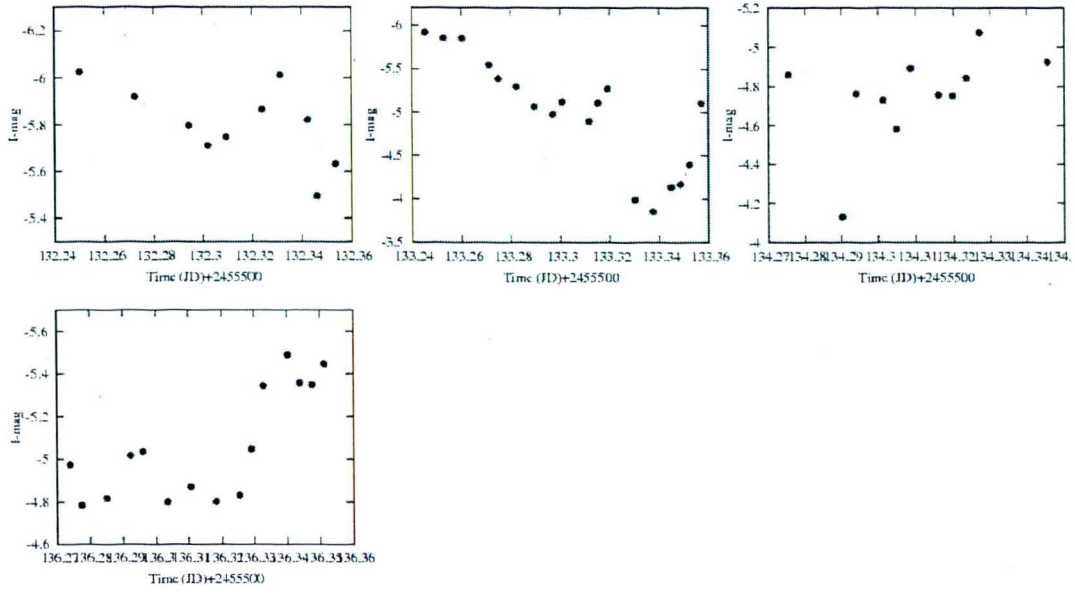


Figure 4.11: A figure showing the light curves of star 6942 from DoPhot reduction outputs in different nights.

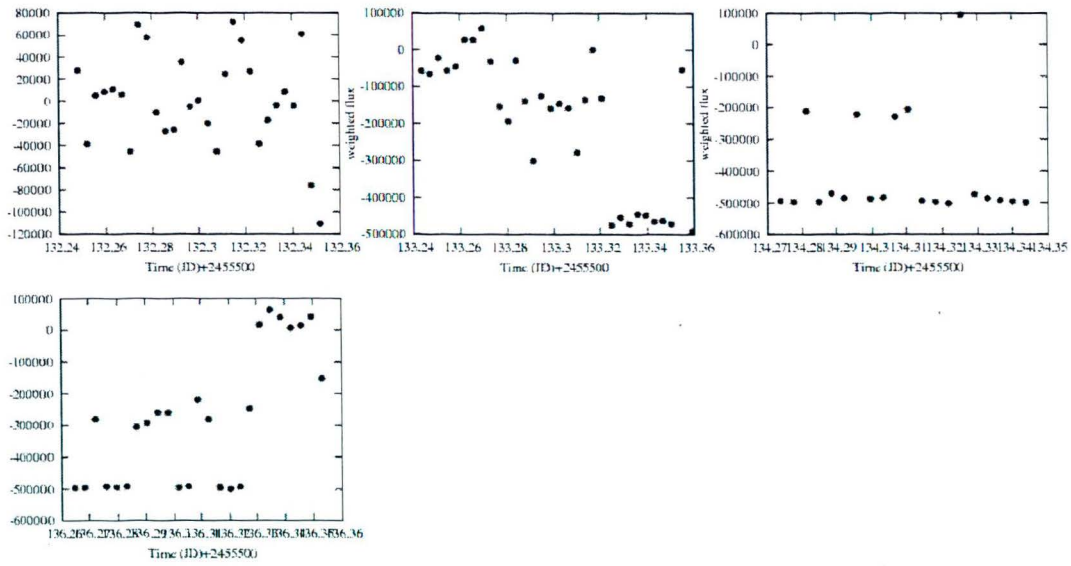


Figure 4.12: A figure showing the light curves of star 844 from ISIS reduction outputs, this is the same star as in Fig.4.11.

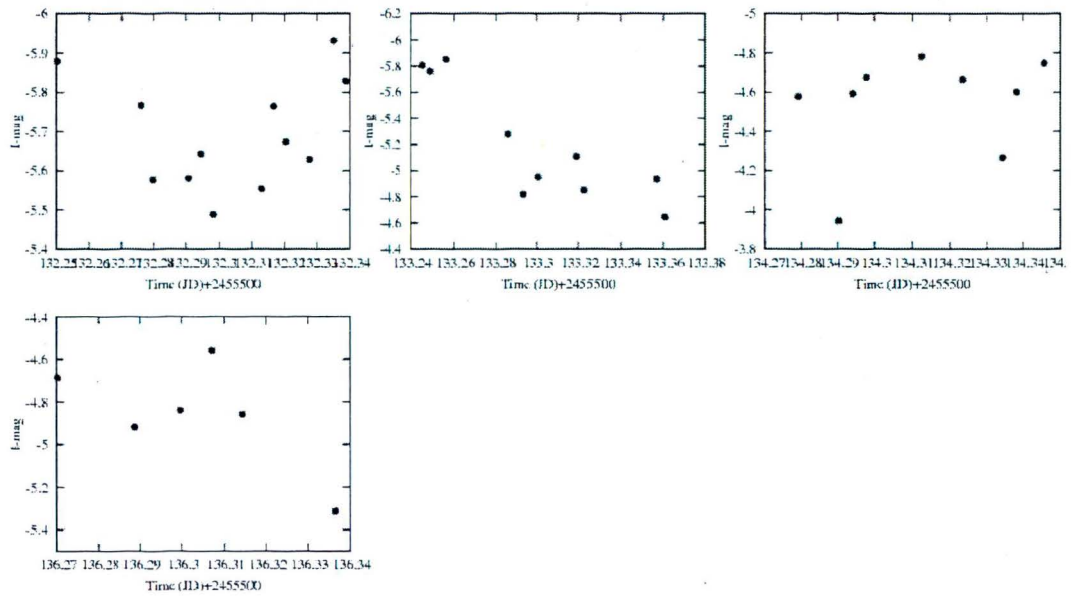


Figure 4.13: A figure showing the light curves of star 6324 from DoPhot reduction outputs in different nights.

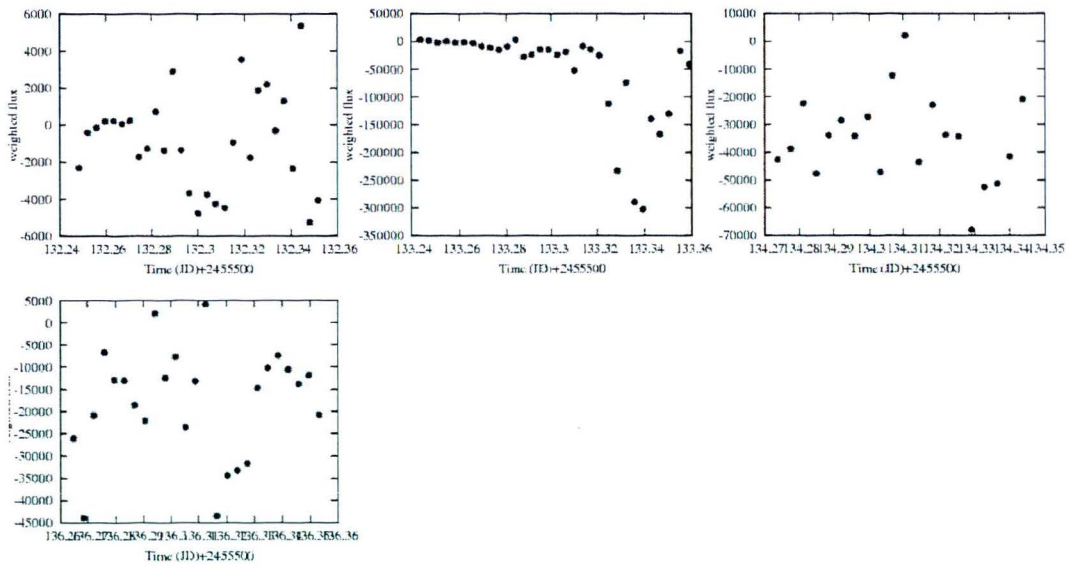


Figure 4.14: A figure showing the light curves of star 1113 from ISIS reduction outputs, this is the same star as in Fig.4.13.

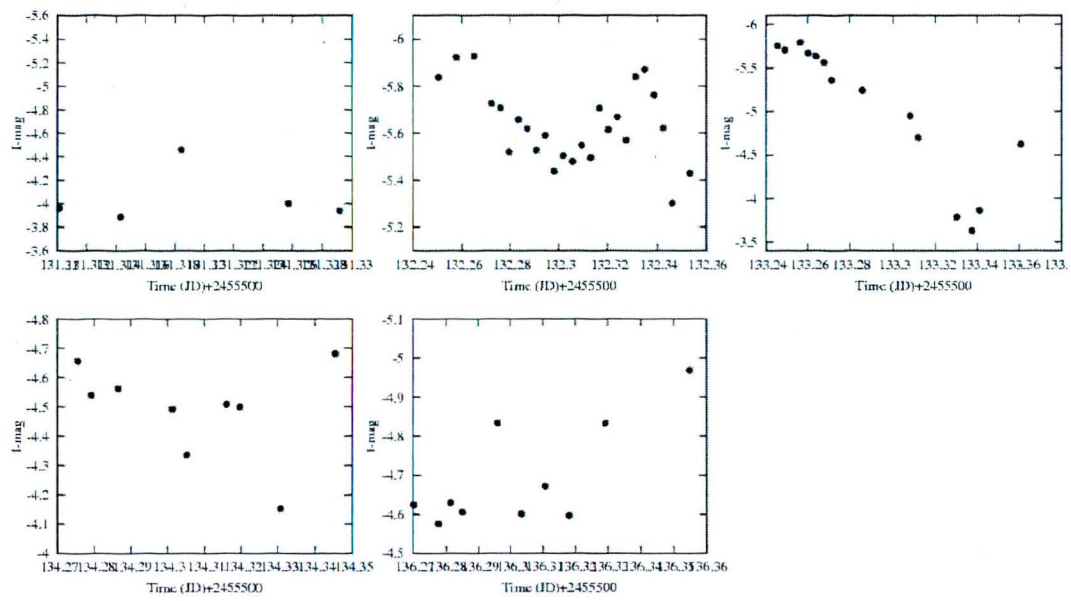


Figure 4.15: A figure showing the light curves of star 6848 from DoPhot reduction outputs in different nights.

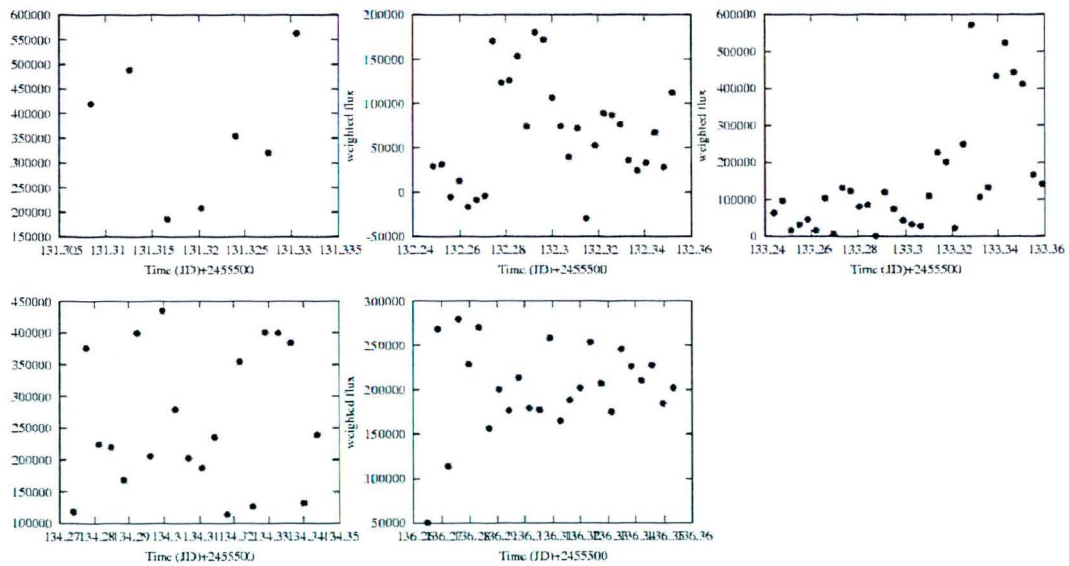


Figure 4.16: A figure showing the light curves of star 1842 from ISIS reduction outputs, this is the same star as in Fig.4.15.

4.2 Colour-magnitude diagrams

Having seen that DoPhot was capable of producing reasonable results, the idea was to use its output to produce colour-magnitude diagrams of the November 2010 data. The magnitudes from Johnson UBVI filters were obtained by applying DoPhot on all the frames of LMC_T1 and LMC_T13 combined. After DoPhot was executed, the `sum` files were separated according to their fields. From these `sum` files, the template `sum` file, per field, was chosen to be used for correcting the stars numbering. After the corrections, the produced `cat` files were separated according to their filters. It should be noted that the correction of the numbering of stars in `sum` files was done with all the `sum` files combined for each field. This was to make sure that all resulting `cat` files have the fixed numbering in which the same number is the same star in all filters. The magnitudes of the stars in `cat` files were calculated as explained in chapter 3.

The first step after having the magnitudes was to apply the comparison stars to do differential photometry. These were selected among the reduced stars. The first ten stars were pre-selected, then after executing the program called `ccdstd`, which returned the magnitude differences and standard deviations, we selected the comparison stars with the smallest standard deviation from the list of ten stars initially selected. These stars were used as comparison stars on the second execution of `ccdstd`, and the output for each star in each filter were saved.

The second reduction step was to get an average magnitude difference for all stars on the frames. This was done with the program called `ccdstats`, which ignored magnitude values greater than 90 (this values represent the missing values due to the reduction program not being able to find the magnitude of such stars). The `ccdstats` also found the standard deviations of both fitted and aperture photometry magnitudes. The calculated magnitudes for each star in all filters were combined into one file. The V-magnitudes were plotted against the colours: (U-B), (U-I), (B-V) and (I-V), for both LMC_T1 and LMC_T13 fields, and these plots are shown on Figs. 4.17, 4.18,

4.19, 4.20 for LMC-T1 and Figs. 4.21, 4.22, 4.23, 4.24, for LMC-T13. The above mentioned analysis programs, i.e. `ccdstd` and `ccdstats`, are Fortran programs kindly provided by Dr. Balona.

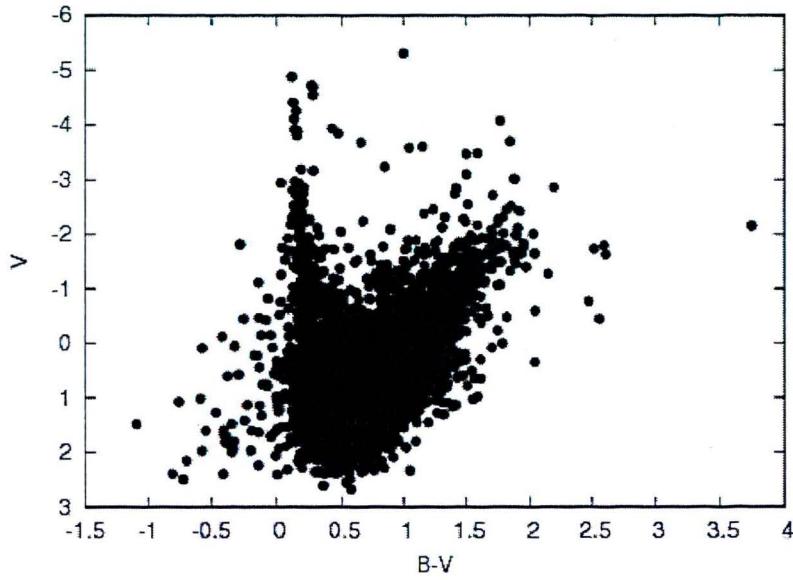


Figure 4.17: A plot showing the relation between V-magnitude and B-V-colour of the LMC-T1 field. The plotted data was reduced using DoPhot.

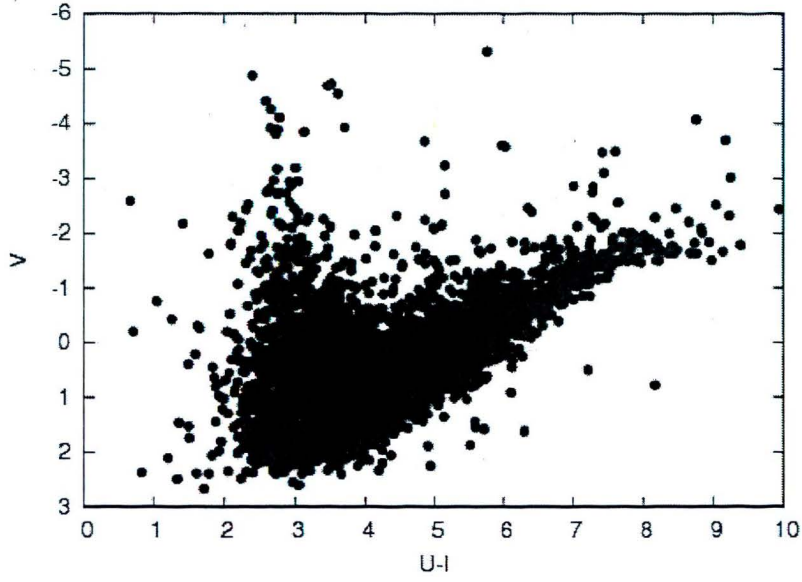


Figure 4.18: A plot showing the relation between V-magnitude and U-I-colour of the LMC.T1 field. The plotted data was reduced using DoPhot.

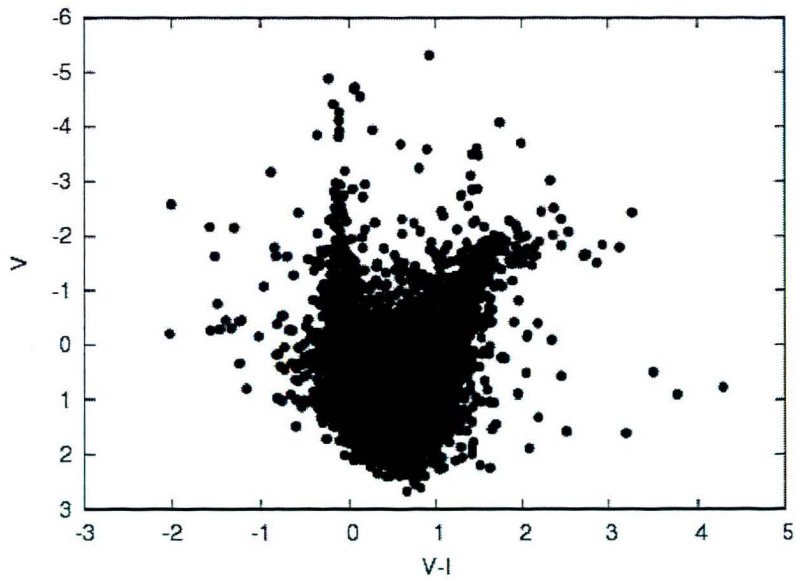


Figure 4.19: A plot showing the relation between V-magnitude and V-I-colour of the LMC.T1 field. The plotted data was reduced using DoPhot.

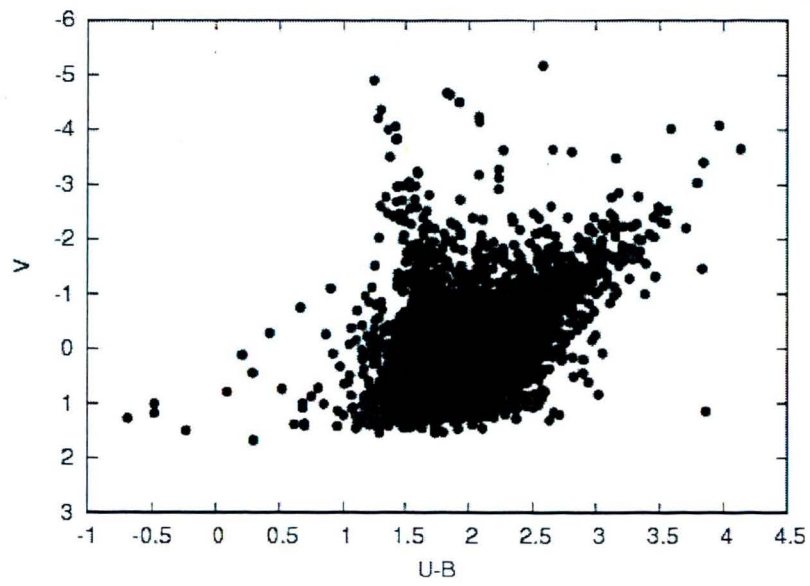


Figure 4.20: A plot showing the relation between V-magnitude and U-B-colour of the LMC.T1 field. The plotted data was reduced using DoPhot.

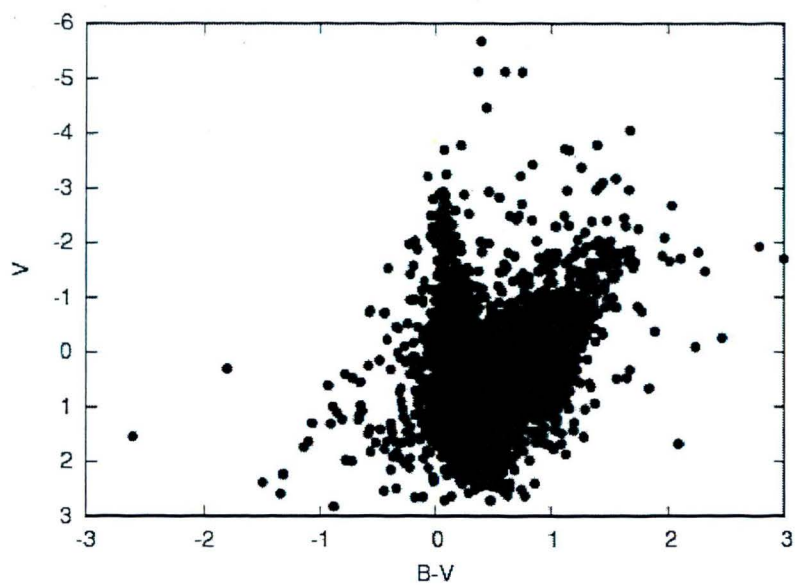


Figure 4.21: A plot showing the relation between V-magnitude and B-V-colour of the LMC_T13 field. The plotted data was reduced using DoPhot.

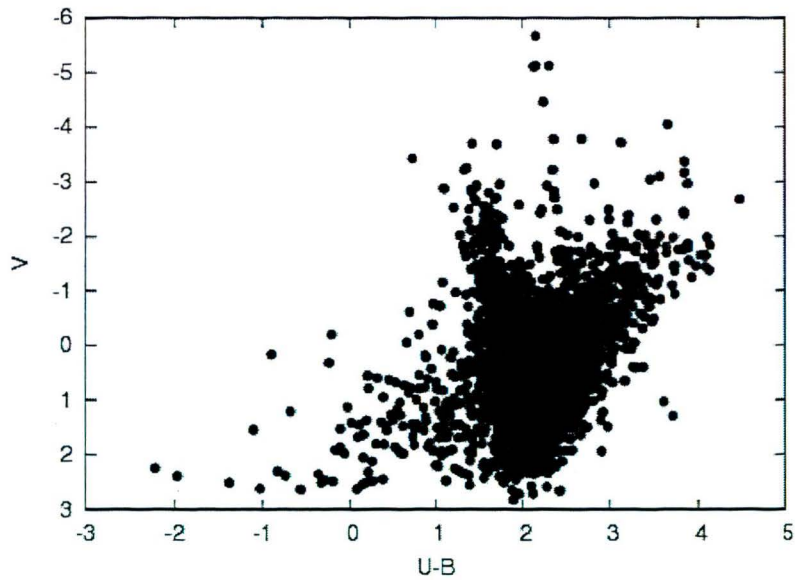


Figure 4.22: A plot showing the relation between V-magnitude and U-B-colour of the LMC_T13 field. The plotted data was reduced using DoPhot.

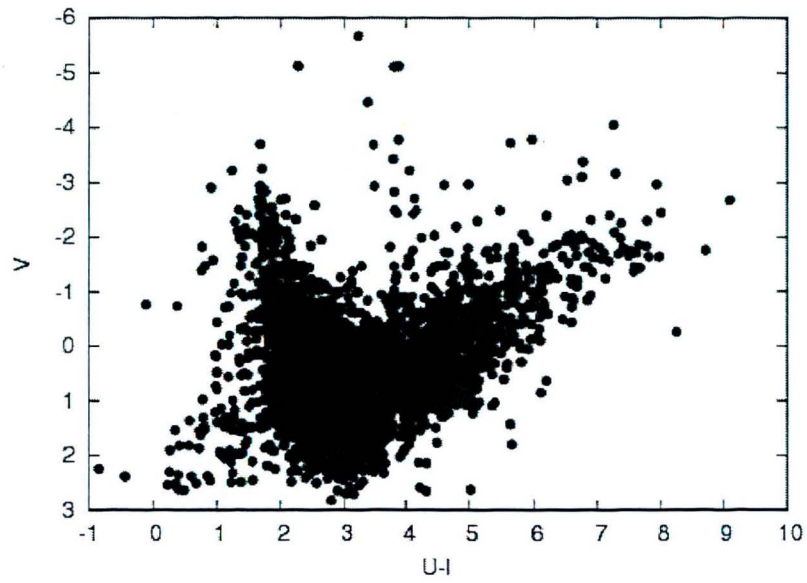


Figure 4.23: A plot showing the relation between V-magnitude and U-I-colour of the LMC_T13 field. The plotted data was reduced using DoPhot.

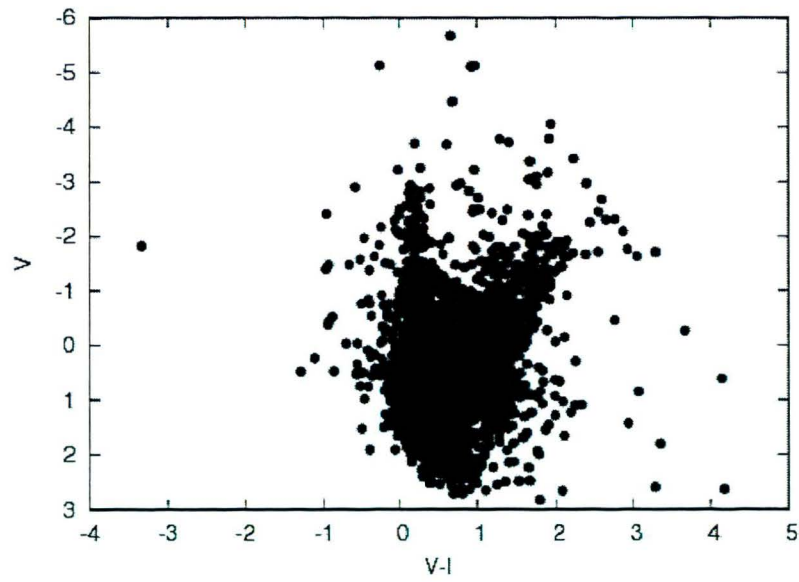


Figure 4.24: A plot showing the relation between V-magnitude and V-I-colour of the LMC_T13 field. The plotted data was reduced using DoPhot.

All the colour-magnitude diagrams show an irregular V-shaped structure, with U-I diagrams of both fields, i.e. Figs. 4.18 and 4.23, having a less steeper right side of the V-shape. The left sides of all the V-shaped structures are very steep and if one thinks of normal the HR-diagram, the steep sides of our diagrams resemble the upper main sequence. The less steep sides consist of the stars which have evolved away from the main sequence. One would expect to find the B-stars on the steep region of the colour-magnitude diagrams. The easy way to find out which stars on these regions are B-stars is to convert the magnitudes to absolute magnitudes in order to produce M_V vs (B-V) HR-diagrams. These require standardized magnitudes and section 4.3 explains the standardization of the magnitudes of stars on the colour-magnitude diagrams above.

It was important to flat-field the raw data on a weekly basis to avoid the effects of atmospheric changes over a long time, meaning that the data taken during a certain week should be flat-fielded with a normalized flat field frame from the flat field frames taken that week. This was realized after normalizing all the flat field frames for the November 2010 campaign and using the normalized flat field frame for flat-fielding. Fig. 4.25 shows two cleaned images from the raw image taken in I-filter. The image on the left was cleaned with a normalized flat-field frame from flat fields that were collected on weekly basis while the image on the right, with vertical stripes, was cleaned with a normalized flat-field frame from all flat-field frames of the November 2010 campaign. Therefore it is clear that the stripes are due to flat-fielding with a normalized flat-field frame of all flat frames of the November 2010 observing campaign.

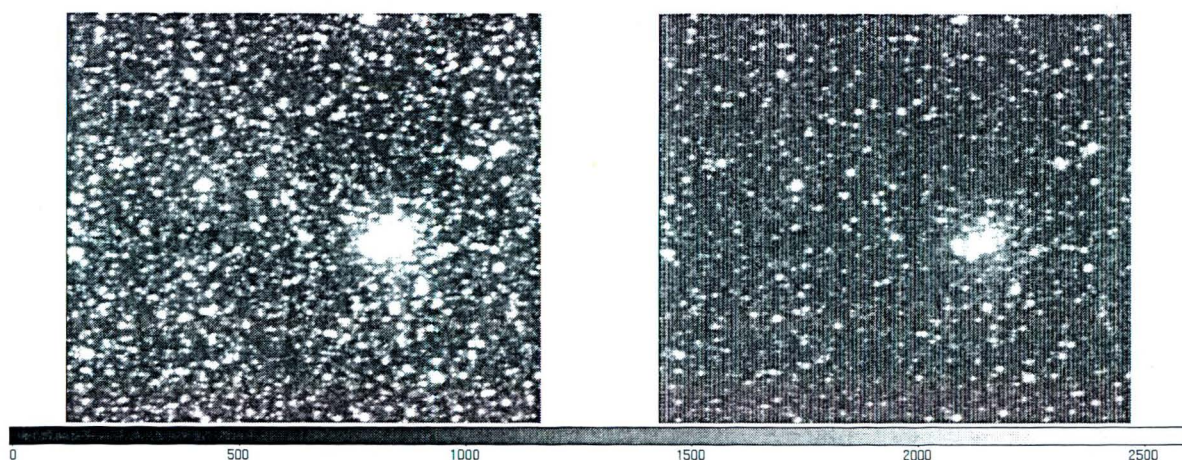


Figure 4.25: A plot showing the effect of flat-fielding raw images with a normalized flat field frame of all flat field frames of an entire observation campaign combined (right image), and raw image flat-fielded with normalized flat field frames of a week (left image).

The effect of the above explained situation was that the results from the two flat-fielding approaches were different. For the case where a normalized flat field frame was produced with all flat frames of November 2010, the correction of the numbers of stars done with a `ccdcatalog` programs was encountering difficulties in identifying stars and as a result ignoring most of the `sum` files. This resulted in obtaining colour-magnitude diagrams with much fewer stars than expected. The colour-magnitude diagrams in Figs. 4.17, 4.18, 4.19, 4.20, 4.21, 4.22, 4.23, 4.24 were produced with the data flat fielded with the flat fields that were collected on a weekly basis.

4.3 Standardizing B,V and I magnitudes

The magnitudes used to produce colour-magnitude diagrams were not standardized and in order to identify β Cephei stars, we needed the diagrams plotted with standardized magnitudes. Magnitudes standardization was done by extracting available data of our fields, i.e. LMC_SC3 and LMC_SC13, from the OGLE II database. We then compared the images of those fields with the CCD images of the template `sum` files used when `ccdcatalog` was executed, these CCD images are `c0010888.fits` for LMC_SC3 and `c0011110.fits` for LMC_SC13. From the OGLE map images and CCD images, a few matching stars were identified and the (x,y) CCD coordinates from CCD images were converted to right ascension (RA) and declination (DEC) coordinates. This was done with a Fortran program provided by Dr. Balona. The program performs linear transformation by first using identified stars from CCD images and OGLE maps to calculate the coefficients of a linear equations:

$$RA = a_1 + a_2x + a_3y \quad (4.1)$$

and

$$DEC = b_1 + b_2x + b_3y \quad (4.2)$$

In order to be certain that we identified correct corresponding stars on CCD images and OGLE II maps, we chose a reference star on the CCD image and same reference star on an OGLE II map. We then calculated the distances between arbitrary corresponding reference stars and other identified stars, using (x,y) coordinates of the stars (in both CCD images and OGLE maps). The ratio of the distances from reference star to any star A and to star B on the CCD image should be approximately the same as the ratio of the distances from the reference star to star A' and to star B' on the OGLE map. This is demonstrated in Fig 4.26 and is expressed mathematically as:

$$\frac{D_{RA}}{D_{RB}} = \frac{d_{R'A'}}{d_{R'B'}} \quad (4.3)$$

where D_{RA} and D_{RB} are the distances on the CCD image and $d_{R'A'}$ and $d_{R'B'}$ are the distances on the OGLE map.

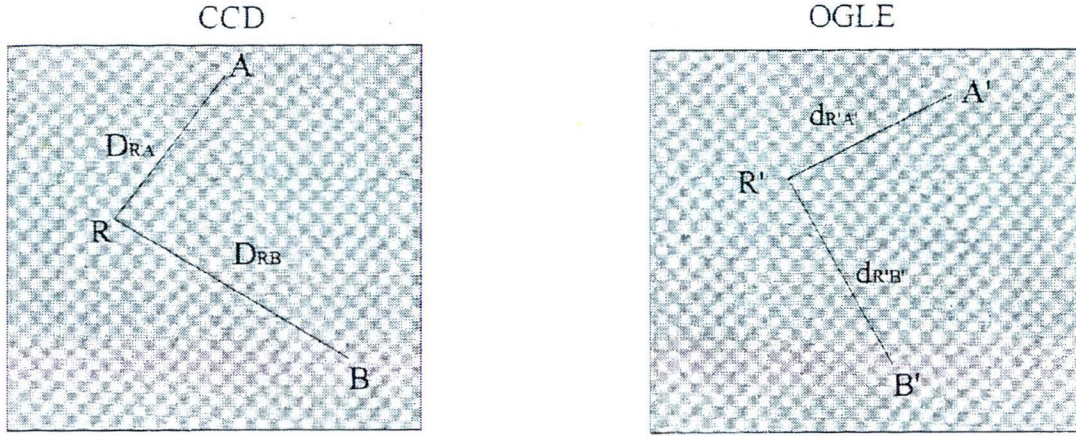


Figure 4.26: Comparison of star patterns in a CCD images and the same pattern in the OGLE maps. The diagrams represent the stars and distances presented in equation (4.3). The left panel diagram represents any stars A, B and an arbitrary reference star R on a the CCD image and right panel diagram represents corresponding stars on the OGLE map.

A python program we wrote to perform the above mentioned calculations is included in section A.2.3 of the Appendix. Tables 4.2 and 4.3 show the resulting distance ratios. The stars we identified in our CCD images and corresponding OGLE II maps are shown in Figs. 4.27 and 4.28.

star number	CCD X	CCD Y	OGLE X	OGLE Y	CCD ratio	OGLE ratio
1	368	318	589.94	7165.25	0.0	0.0
2	385	362	622.56	7102.04	1.0	1.0
3	53	466	141.73	6899.60	0.1355	0.1365
4	130	420	251.24	6981.93	0.1822	0.1847
5	220	246	360.89	7251.26	0.2866	0.2901
6	219	219	365.49	7185.23	0.3115	0.3157
7	174	287	298.48	7184.64	0.2401	0.2435
8	370	434	609.32	6993.73	0.4066	0.4121

Table 4.2: A table showing the coordinates and distance ratios of the stars identified in the LMC_T1 field and its corresponding OGLE II map. The last two columns show the distance ratios presented in equation (4.3).

star number	CCD X	CCD Y	OGLE X	OGLE Y	CCD ratio	OGLE ratio
1	479	333	1922	1801	0.0	0.0
2	442	320	1865	1817	1.0	1.0
3	393	179	1775	2020	0.2223	0.2245
4	363	161	1728	2044	0.1890	0.1904
5	263	236	1589	1918	0.1675	0.1677
6	216	307	1528	1808	0.1484	0.1502
7	157	293	1437	1821	0.1209	0.1219
8	474	203	1899	1995	0.3014	0.3030

Table 4.3: A table showing the coordinates and distance ratios of the stars identified in the LMC_T13 field and its corresponding OGLE II map. The last two columns show the distance ratios presented in equation (4.3).

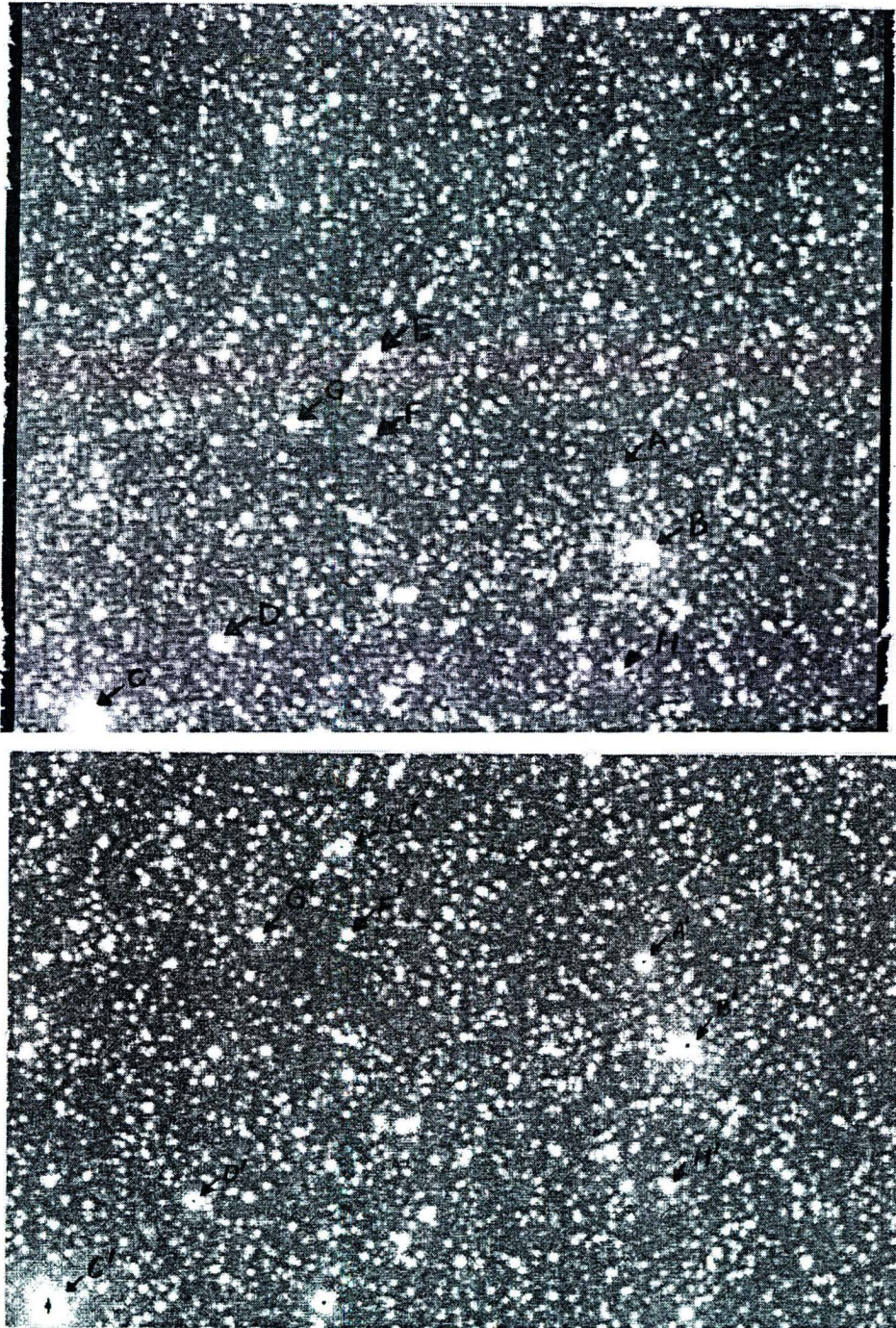


Figure 4.27: The top image shows stars in the CCD frame we observed in the LMC.T1 field. The bottom image shows the same group of stars in the OGLE II map identified for calculating coefficients in equations (4.1) and (4.2). Identified stars are pointed with arrows and marked with letters.

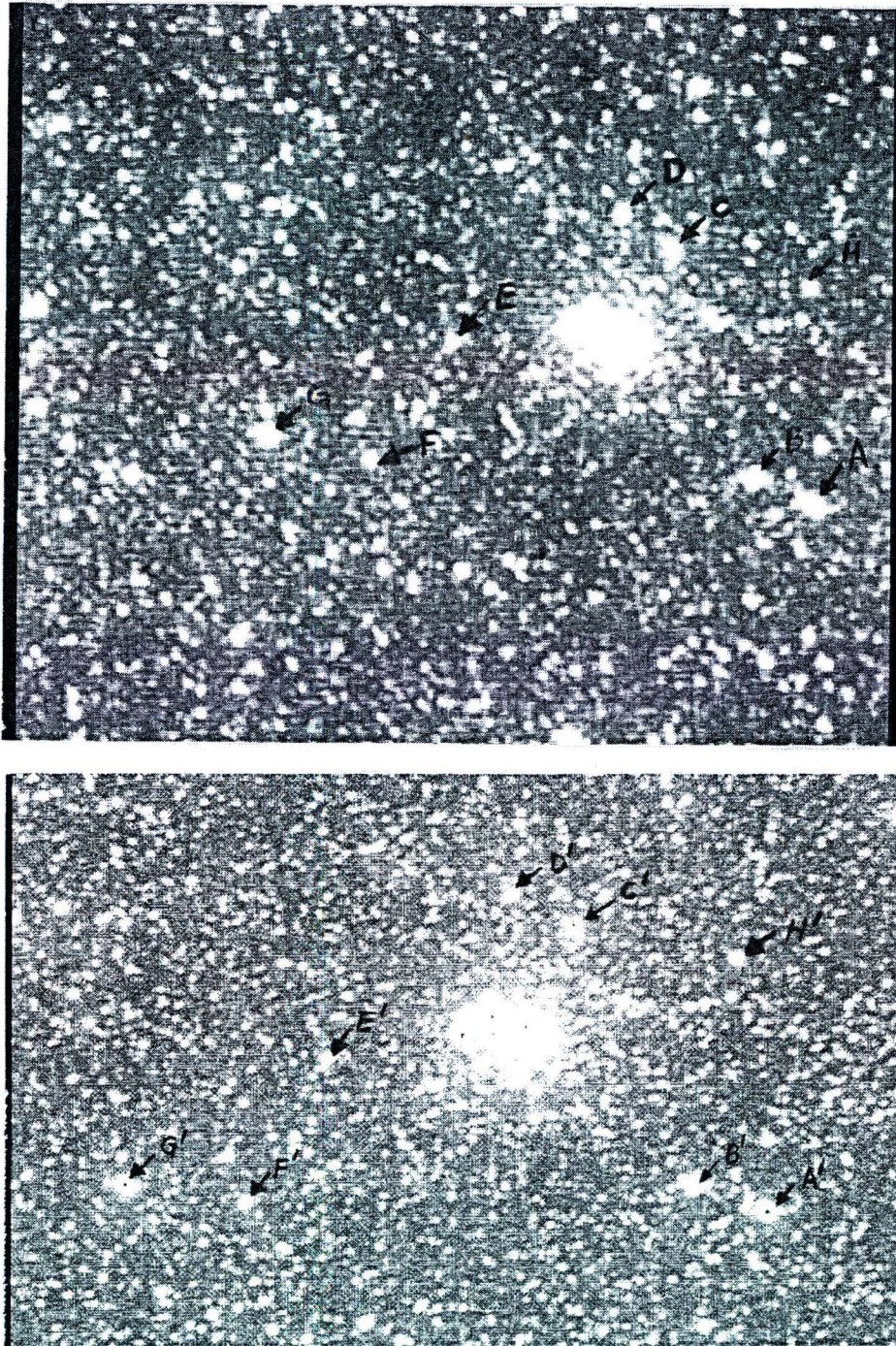


Figure 4.28: The top image shows stars in the CCD frame we observed in the LMC.T13 field. The bottom image shows the same group of stars in the OGLE II map identified for calculating coefficients in equations (4.1) and (4.2). Identified stars are pointed with arrows and marked with letters.

After verifying identified stars, a Fortran program provided by Dr. Balona calculated the values of the coefficients a_i and b_i by means of a multivariable least squares fit. These coefficients were used to calculate RA and DEC of all the stars in our fields. We matched RA and DEC coordinates of our stars with the coordinates of the stars observed in OGLE II to within 5 arc-seconds. This allowed us to have the true B, V and I magnitudes for matching stars. A python program used for matching coordinates is attached in section A.2 of the Appendix. The idea for matching coordinates was to use the magnitudes of OGLE II stars that matched with stars on our fields to standardize the magnitudes on our colour-magnitude diagrams, since we did not measure standard stars during our observing run. We could not standardize U magnitudes because of unavailability of U magnitudes for OGLE II stars.

To standardize our B, V and I magnitudes, we first calculated magnitude differences between OGLE II magnitudes and our magnitude for each of the matched stars. That is:

$$M_D = M_{\text{OGLEII}} - M_{\text{SAAO}} \quad (4.4)$$

where M_D represents magnitudes difference, M_{OGLEII} represents OGLE II magnitudes and M_{SAAO} represents magnitudes of stars we observed in Sutherland. Among the matched stars with magnitude differences calculated, we selected the 30 brightest stars and calculated the standard deviation of M_D and average value ($\bar{M}_D$). We further calculated the difference between $\bar{M}_D$ and M_D and stars with this difference greater than three times the standard deviation were eliminated from the list of 30 brightest stars. This procedure of eliminating stars by calculating standard deviation and average values, and comparing the difference of $\bar{M}_D$ and M_D with three times standard deviation was continued until there were 10 brightest stars left with smallest deviation and these stars are listed in Tables 4.4, 4.5, 4.6 4.7, 4.8 and 4.9. The above mentioned analysis was done for all B, V and I magnitudes.

B filter					
OGLE ID	B _{SAAO}	B _{OGLE}	B _{corrected}	RA	DEC
101996	-0.010	15.274	15.723	05 27 46.8	-69 29 32.5
108072	0.0	15.384	15.733	05 27 31.6	-69 26 05.4
220005	0.0	15.878	15.733	05 28 27.2	-69 24 31.0
220075	0.011	16.425	15.744	05 28 29.3	-69 26 01.3
108019	0.223	16.433	15.956	05 27 51.1	-69 26 46.9
102225	1.157	16.574	16.890	05 33 54.2	-69 28 26.2
102188	1.260	16.686	16.993	05 27 39.0	-69 29 09.3
214147	0.883	16.703	16.616	05 28 12.3	-69 28 10.7
102542	1.761	17.253	17.494	05 28 04.9	-69 28 50.2
102396	1.297	17.304	17.030	05 28 03.3	-69 29 43.4

Table 4.4: A table showing a list of OGLE II stars used for determining the standardizing factors for the B magnitudes of stars in LMC-T1. Column 2 contain magnitudes of stars we observed at Sutherland, column 3 contain the magnitudes from OGLE II observations and column 4 are corrected magnitudes.

V filter					
OGLE ID	V_{SAAO}	V_{OGLE}	$V_{\text{corrected}}$	RA	DEC
108019	-0.841	15.046	15.400	05 27 51.1	-69 26 46.9
220005	-0.009	16.030	16.232	05 28 27.2	-69 24 31.0
102002	0.0	16.108	16.241	05 28 01.1	-69 28 15.0
108026	-0.212	16.404	16.029	05 27 51.6	-69 24 46.8
108103	0.009	16.481	16.250	05 27 52.5	-69 24 11.0
102225	0.469	16.569	16.710	05 27 50.6	-69 28 26.2
220122	0.446	16.597	16.687	05 28 23.9	-69 25 01.8
220075	0.0	16.603	16.241	05 28 29.3	-69 26 01.3
214194	0.682	16.794	16.923	05 28 11.2	-69 27 25.6
214147	0.498	16.821	16.739	05 28 12.3	-69 28 10.7

Table 4.5: A table showing a list of OGLE II stars used for determining the standardizing factors for the V magnitudes of stars in LMC-T1. Column 2 contain magnitudes of stars we observed at Sutherland, column 3 contain the magnitudes from OGLE II observations and column 4 are corrected magnitudes.

I filter					
OGLE ID	I _{SAAO}	I _{OGLE}	I _{corrected}	RA	DEC
108026	-1.483	13.964	13.890	05 27 51.6	-69 24 46.8
219949	-1.608	13.995	13.763	05 28 26.6	-69 24 17.5
219961	0.0	15.017	15.371	05 28 12.9	-69 24 48.1
102037	-0.024	15.252	15.347	05 27 48.2	-69 27 32.9
214022	0.076	15.477	15.447	05 28 30.7	-69 27 33.5
108039	0.0	15.504	15.371	05 27 46.1	-69 26 18.1
213957	0.024	15.527	15.395	05 28 08.7	-69 27 55.1
102102	0.010	15.572	15.381	05 28 34.4	-69 27 52.5
102093	0.538	15.608	15.954	05 27 55.0	-69 28 21.6

Table 4.6: A table showing a list of OGLE II stars used for determining the standardizing factor for I magnitudes of stars in LMC_T1. Column 2 contains magnitudes of stars we observed at Sutherland, column 3 contains the magnitudes from OGLE II observations and column 4 are corrected magnitudes.

B filter					
OGLE ID	B _{SAAO}	B _{OGLE}	B _{corrected}	RA	DEC
218738	-0.011	15.958	16.250	05 07 10.4	-68 58 26.2
218744	0.0	15.973	16.261	05 07 07.4	-68 58 17.6
218742	-0.616	15.992	15.645	05 07 08.4	-68 58 21.8
156150	-0.197	16.066	16.064	05 06 43.1	-68 56 18.8
218737	-0.144	16.112	16.117	05 07 06.2	-68 58 32.1
218857	0.0	16.366	16.261	05 07 06.1	-68 58 20.6
147502	0.012	16.465	16.273	05 06 34.7	-69 01 40.7
218817	0.494	16.465	16.755	05 07 02.1	-68 58 58.7
218858	0.0	16.494	16.261	05 07 17.5	-68 58 19.7

Table 4.7: A table showing a list of OGLE II stars used for determining the standardizing factors for B magnitudes of stars in LMC_T13. Column 2 contain magnitudes of stars we observed at Sutherland, column 3 contain the magnitudes from OGLE II observations and column 4 are corrected magnitudes.

V filter					
OGLE ID	V_{SAAO}	V_{OGLE}	$V_{\text{corrected}}$	RA	DEC
218738	0.0	15.849	16.181	05 07 10.4	-68 58 26.2
218696	-0.215	15.949	15.966	05 07 18.9	-68 59 08.4
218744	0.01	15.981	16.191	05 07 07.4	-68 58 17.6
218737	-0.154	16.062	16.027	05 07 06.2	-68 58 32.1
151841	0.0	16.174	16.181	05 06 36.7	-68 59 30.4
218857	-0.019	16.301	16.162	05 07 06.1	-68 58 20.6
218774	-0.01	16.345	16.171	05 07 03.4	-69 00 22.1
218695	0.020	16.394	16.201	05 07 18.7	-68 59 12.6
218817	0.266	16.472	16.447	05 07 02.1	-68 58 58.7

Table 4.8: A table showing a list of OGLE II stars used for determining the standardizing factors for V magnitudes of stars in LMC_T13. Column 2 contain magnitudes of stars we observed at Sutherland, column 3 contain the magnitudes from OGLE II observations and column 4 are corrected magnitudes.

I filter					
OGLE ID	I _{SAAO}	I _{OGLE}	I _{corrected}	RA	DEC
218699	-0.538	14.495	14.574	05 07 09.5	-68 58 49.3
218696	-0.929	14.503	14.183	05 07 18.9	-68 59 08.4
147371	-0.019	14.737	15.093	05 06 43.4	-69 01 34.4
218695	0.0	14.845	15.112	05 07 18.7	-68 59 12.6
218685	-0.040	15.122	15.072	05 07 13.8	-69 00 50.0
214145	0.0	15.140	15.112	05 07 17.9	-69 01 38.7
223869	0.020	15.159	15.131	05 07 08.5	-68 57 16.4
223875	0.0	15.165	15.112	05 07 20.6	-68 56 28.5
223873	0.041	15.279	15.152	05 07 09.0	-68 56 56.7
151906	0.075	15.280	15.187	05 06 34.3	-68 57 41.7

Table 4.9: A table showing a list of OGLE II stars used for determining the standardizing factor for I magnitudes of stars in LMC.T13. Column 2 contains magnitudes of stars we observed at Sutherland, column 3 contains the magnitudes from OGLE II observations and column 4 are corrected magnitudes.

The average difference values ($\bar{M}_D$'s) for B, V and I magnitudes of the 10 remaining stars in Tables 4.4, 4.5, 4.6, 4.7, 4.8 and 4.9 were used as the standardizing factor by adding them to our respective unstandardized magnitude values. Calculated average difference values and their standard deviations in B, V and I filters are shown in Table 4.10.

Field (LMC_)	Filters	standardizing factor (mag.)	standard deviation (mag.)
T1	B	15.733	0.390
T1	V	16.241	0.249
T1	I	15.371	0.220
T13	B	16.261	0.244
T13	V	16.181	0.174
T13	I	15.112	0.194

Table 4.10: A table showing standardizing magnitude factors (average difference) and their standard deviations obtained for correcting our B, V and I magnitudes in Figs. 4.17, 4.18, 4.19, 4.20, 4.21, 4.22, 4.23, 4.24.

From the standardized B,V and I magnitudes, we plotted two colour-magnitude diagrams: V vs B-V and V vs V-I, these diagrams are shown in Fig. 4.29.

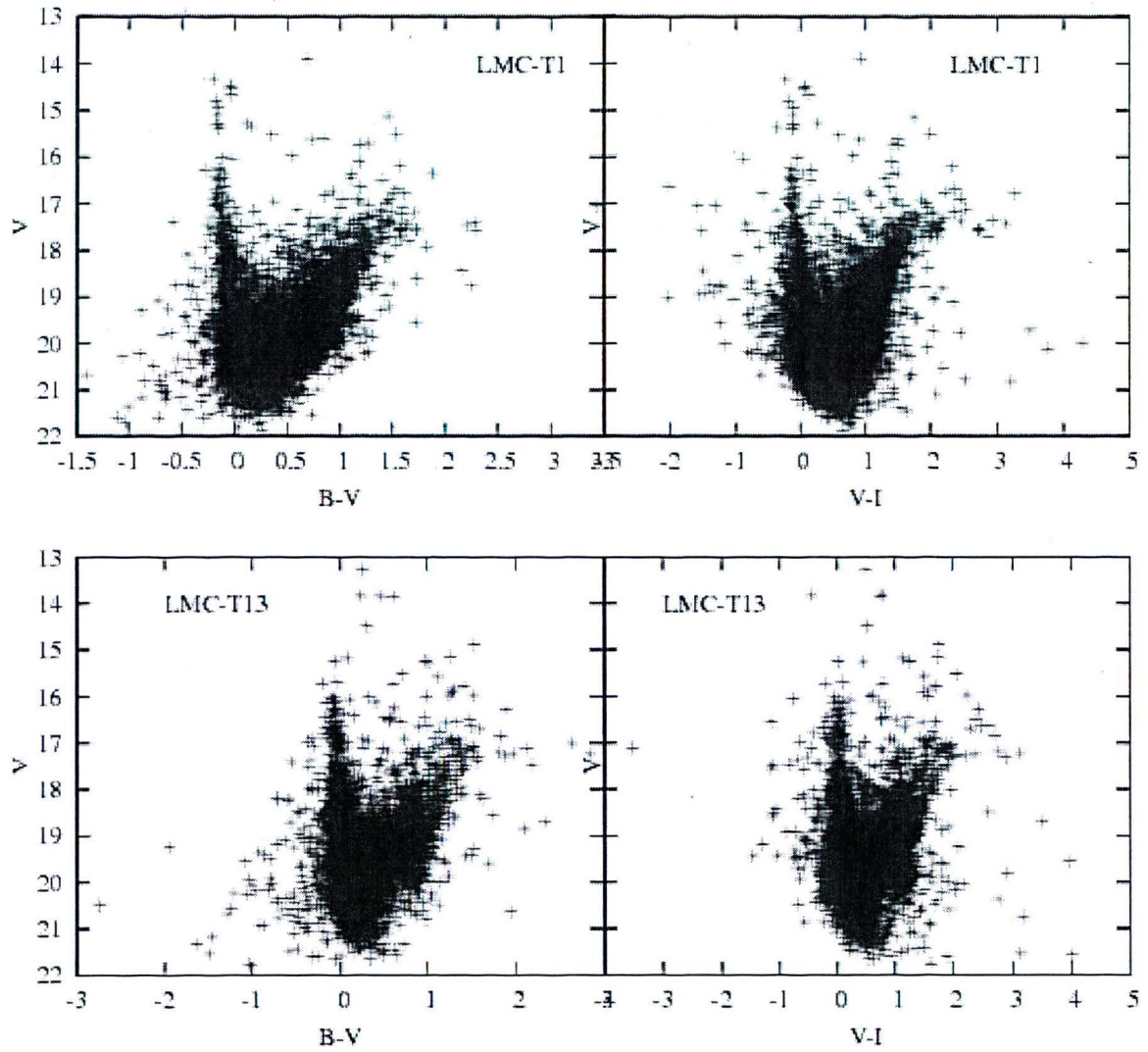


Figure 4.29: A figure showing colour-magnitude diagrams of LMC-T1 (upper panels) and the LMC-T13 (lower panels) from standardized magnitudes. It can be seen that these diagrams are consistent with those plotted with unstandardized magnitudes on Figs. 4.17, 4.19, 4.21 and 4.24.

The standardized colour-magnitude diagrams in Fig. 4.29 look similar to those plotted with an unstandardized magnitudes (see Figs. 4.17, 4.19, 4.21 and 4.24). This validates the standardization approach we applied on our magnitudes as we do not expect standardization to affect the shape of the distribution of stars. The shapes of the above colour-magnitude diagrams also resemble those of Pigulski & Kolaczowski (2002) and Kolaczowski et al. (2004) (see Figs. 1.15 and 1.16). Comparing our dia-

grams with theirs it can be seen that β Cephei stars are located on the left side of the V-shaped structure of the diagrams.

Having standardized apparent magnitudes, absolute magnitude values (M_B , M_V and M_I) were calculated using the distance modulus equation:

$$m - M = 5 \log_{10}(d) - 5 \quad (4.5)$$

The letter m in equation (4.5) represents apparent magnitude, M is an absolute magnitude and d is the distance in parsec. We used a distance modulus of 18.52 mag (Bonanos et al. 2011) for the LMC.

For the purpose of identifying β Cephei stars, we plotted the absolute magnitude in the V-filter against (B-V) colour. The absolute magnitude (M_V) range for β Cephei stars was determined from the fact that the hottest β Cephei stars have $M_V = -5.1$ mag and the coolest have $M_V = -2.1$ mag. The colour (B-V) range of β Cephei stars was determined from their temperature range of 22570 K for B3V stars to 30200 K for O9III stars¹, and the polynomial:

$$\log T_{\text{eff}} = a + b(B - V) + c(B - V)^2 + \dots \quad (4.6)$$

taken from Torres (2010), where T_{eff} is an effective temperature, (B-V) is the colour. The coefficients a , b , c ,...etc were also taken from Torres (2010) and are listed in Table 4.11.

¹<http://isthe.com/chongo/tech/astro/HR-temp-mass-table-bymass.html>

Coefficients	Supergiants	Main-Sequence stars
		Subgiants, Giants
a	4.012559732366214	3.979145106714099
b	-1.055043117465989	-0.654992268598245
c	2.133394538571825	1.740690042385095
d	-2.459769794654992	-4.608815154057166
e	1.349423943497744	6.792599779944473
f	-0.283942579112032	-5.396909891322525
g	...	2.192970376522490
h	...	-0.359495739295671

Table 4.11: A table of coefficients for the polynomial used to calculate the effective temperature as a function of colour in equation (4.6) (taken from Torres 2010).

These coefficients were initially calculated by Flower (1996) and corrected by Torres (2010). Flower (1996) fitted a polynomial on the plot of $\log T_{\text{eff}}$ vs B-V of 335 stars to determine the values of above mentioned polynomial coefficients. On the data for Supergiant stars, a fifth order polynomial was fitted and the seventh order polynomial was fitted on the data of Main-sequence, Subgiant and Giant stars.

The plots of the absolute magnitude versus colour index for LMC_T1 and LMC_T13 are shown in Fig. 4.30.

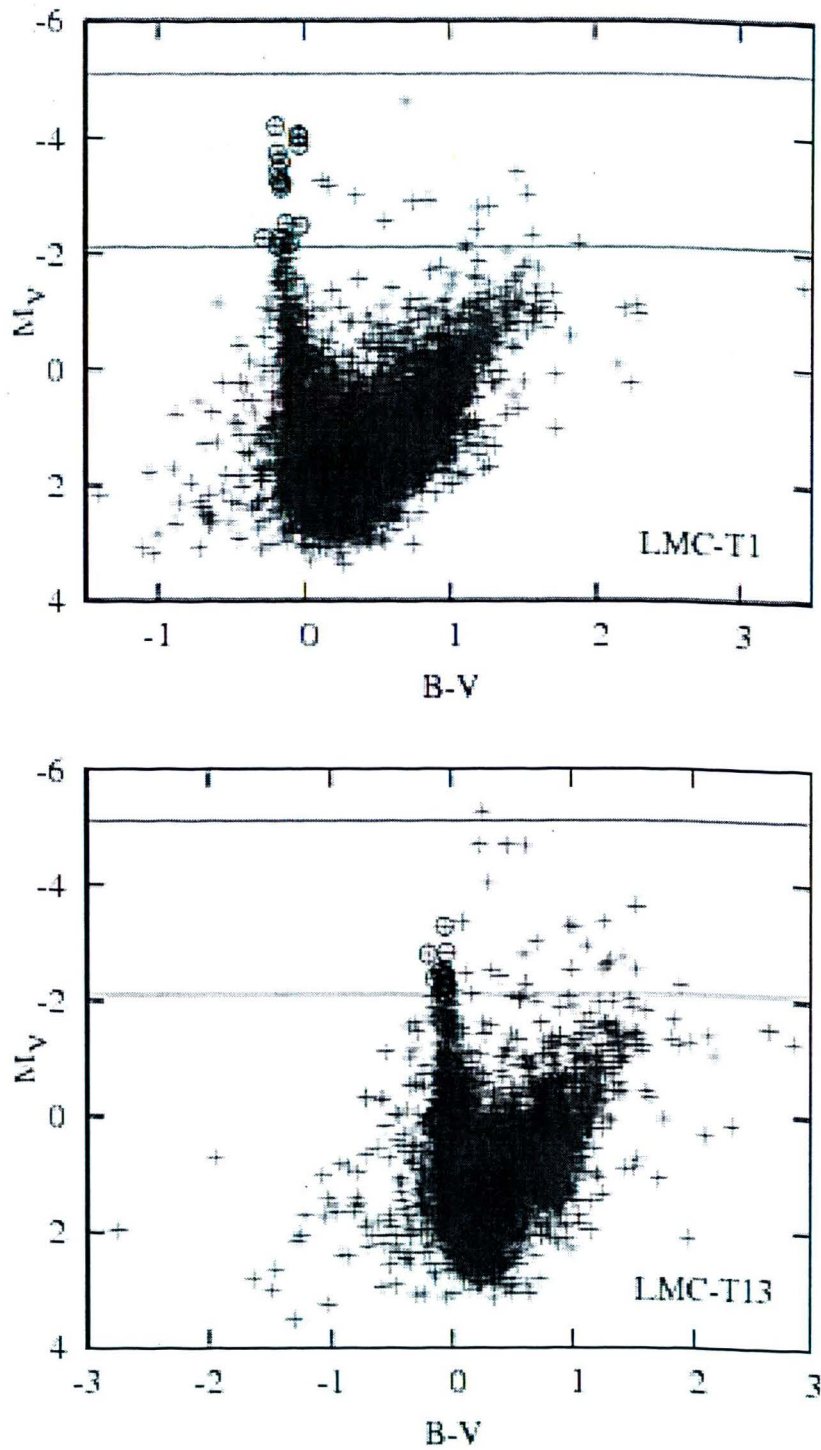


Figure 4.30: Figures showing absolute magnitude (M_V) against the colour ($B-V$) for the LMC-T1 (top panel) and the LMC-T13 (bottom panel) fields. Horizontal lines mark M_V range of β Cephei stars. Suspected β Cephei stars are marked with crosses inside circles.

From Fig. 4.30 we extracted the stars which are within the colour and absolute magnitude ranges of β Cephei stars. These stars, marked by crossed circles in Fig. 4.30, are presented in Table 4.12 for LMC_T1 and Table 4.13 for LMC_T13.

ID	X	Y	RA	DEC	V	B	I
1	158.70	112.68	05 27 46.3	-69 25 01.6	14.3273	14.1325	14.5646
2	370.0	433.31	05 28 15.3	-69 28 46.2	14.5166	14.4848	14.4650
5	226.92	386.82	05 27 57.1	-69 27 45.8	15.2884	15.1159	15.4062
9	85.82	219.0	05 27 38.9	-69 26 10.6	15.3197	15.1705	15.4378
15	376.51	95.86	05 28 11.6	-69 24 39.6	16.0182	15.8952	16.0717
21	252.24	409.48	05 27 52.6	-69 24 11.8	16.2478	16.0780	16.4042
22	401.13	407.64	05 28 15.5	-69 25 03.1	16.2726	16.1331	16.3767
6710	53.47	173.13	05 27 25.6	-69 25 50.9	14.9469	14.7849	15.0649
6713	13.26	206.90	05 27 27.9	-69 26 10.1	15.0965	14.9204	15.2208
18	4.35	205.76	05 27 30.4	-69 26 06.4	16.2642	15.9894	16.0916
3	233.33	448.78	05 27 58.5	-69 28 23.4	14.6559	14.6242	14.5327

Table 4.12: A table showing candidate β Cephei stars we obtained from the LMC_T1 field. The second and third columns show the X and Y positions of the stars on our CCD images. Their right ascension and declination coordinates are shown in columns 4 and 5. Also shown on this table are V, B and I magnitudes.

ID	X	Y	RA	DEC	V	B	I
4	358.97	331.06	05 07 09.3	-68 58 27.0	15.2463	15.1858	15.2252
6	457.09	242.16	05 07 19.5	-68 58 24.0	15.6997	15.6578	15.6033
7	449.41	140.62	05 07 18.1	-68 57 21.6	15.7323	15.5286	15.9197
16	342.32	212.78	05 06 56.6	-68 59 29.1	16.2615	16.1910	16.2895
8717	277.42	425.09	05 07 01.8	-69 01 27.8	16.0095	15.9406	16.0423
8822	497.13	456.94	05 07 27.7	-69 01 44.4	16.0901	16.0428	16.0617
20	210.04	248.41	05 06 54.1	-68 59 41.6	16.3016	16.2297	16.2926

Table 4.13: A table showing candidate β Cephei stars we obtained from the LMC-T13 field. The second and third columns show the X and Y positions of the stars on our CCD images. Their right ascension and declination coordinates are shown in columns 4 and 5. Also shown on this table are V, B and I magnitudes.

In order to make a conclusive proof of detecting β Cephei stars in LMC, stars in Tables 4.12 and 4.13 should be examined for periodic variations typical for β Cephei stars. The periodic variations examination might require more observations of individual candidate β Cephei stars, this will improve the possibility of detecting the frequency peaks and reduce the noise level on the periodograms.

Chapter 5

Conclusion

We have searched for possible β Cephei stars in the OGLE II fields. This was done by performing period analysis of light curves we extracted from the OGLE II database. Period analysis was done by making periodograms of extracted light curves. The frequency spectra of 100 targets with smallest FAP were visually inspected for the highest amplitude peaks within the pulsation frequency range of β Cephei stars. We found 16 candidate β Cephei stars and three of the 16 were observed.

The obtained data allowed us to compare the outputs of two data reduction programs, i.e. DoPhot and ISIS. This was done after realizing that the initial program, DUPHOT, was not suitable for reducing crowded field images. It was shown that DoPhot is capable of producing reasonable photometry of crowded field CCD images. However, photometry outputs from DoPhot and ISIS could not be compared directly because of different magnitude units returned by these two data reduction programs.

The PSF fitted magnitudes from DoPhot reductions were used to produce colour-magnitude diagrams of the fields: LMC_T1 and LMC_T13. These were done to identify potential β Cephei stars on the fields. We were able to detect 18 potential β Cephei candidates. We proceeded further by standardizing our magnitudes. Magnitudes standardization was done by matching our data with OGLE II data of the same fields. The magnitudes of the matched stars from OGLE II were used to find standardizing factors

which were added to our magnitudes. The two colour-magnitude diagrams for each of our fields were presented. The similarity of these plotted diagrams with the diagrams plotted with unstandardized magnitudes validated our standardizing approach.

From our standardized magnitudes we calculated absolute magnitudes and made M_V vs $(B-V)$ diagrams for stars in LMC_T1 and LMC_T13 fields. We were able to show that indeed some of our observed candidates are within the absolute magnitude and colour limits of β Cephei stars.

5.1 Future work

To further this work, it is important to Fourier analyze the light curves of all candidate β Cephei stars in Tables 4.12 and 4.13 presented in chapter 4, such analysis will be to search for pulsation frequencies. This further work might require more observations of individual stars identified. The observations on one colour will suffice to do frequency analysis. The few data points we obtained from observations of this thesis can be used to get preliminary results, however, the periodograms will have peaks with large widths and thus it will not be possible to resolve the frequencies.

Bibliography

1. Aerts, C., De Cat, P., 2003, *Space Science Reviews* 105, 453
2. Aert, C., Christensen-Dalsgaard, J., Kurtz, D.W., 2009, *Asteroseismology*, 1st Edition, Springer
3. Alard, C., Lupton, R.H., 1998, *ApJ* 503, 325
4. Asplund, M., Grevesse, N., Sauval, A.J., Allende Prieto, C., Blomme, R., 2005, *A&A* 431, 693
5. Balona, L.A., 1977, *Mem. R. astr. Soc.* 84, 101
6. Balona, L.A., 1983, *Mem. R. astr. Soc.* 203, 1041
7. Balona, L.A., 1995, *MNRAS* 277, 1547
8. Bonanos, A.Z., Castro, N., Macri, L.M., Kudritzki, R.-P., 2011, *ApJ* 729L, 9B
9. Carroll, B.W., Ostlie, D.A., 2007, *An Introduction to Modern Astrophysics*, Second Edition, Pearson Addison Wesley
10. Christensen-Dalsgaard, J., 1999, *Astrophys. Space Science* 261, 1
11. Cox, A.N., Morgan, S.M., Rogers, F.J., Iglesias, C.A., 1992, *ApJ* 393, 272
12. Dahn, C.C. et al, 2002, *ApJ* 124, 1170
13. Daszyńska-Daszkiewicz, J., Niemczura, E., 2003, *AdSpR* 31, 387
14. Diago, P.D. et al, 2009, *A&A* 506, 125
15. Drobek, D., Pigulski, A., Shobbrook, R.R., Narwid, A., 2010, *AN* 331, 1077
16. Dziembowski, W.A., Moskalik, P., Pamyatnykh, A.A., 1993, *MNRAS* 265, 588
17. Flower, P.J., 1996, *ApJ* 469, 355
18. Frost, E.B., 1902, *ApJ* 15, 340

19. Frost, E.B., 1906, ApJ 24, 259
20. Grevesse, N., Noels, A., 1993, in Hauck B. Paltani S.R.D., eds, La Formation des Éléments Chimiques. AVCP, Lausanne p. 205
21. Handler, G., Meingast, S., 2011, A&A 533, A70
22. Hill, G., 1967, ApJS 14, 263
23. Jakate, S.M., 1979, AJ 84, 552
24. Jerzykiewicz, M., 1972, PASP 84, 718
25. Jerzykiewicz, M., Handler, G., Shobbrook, R.R., Pigulski, A., Medupe, R., Mokgwetsi, T., Tlhagwane, P., Rodriguez, E., 2005, MNRAS 360, 619
26. Kaler, J.B., 1997, Stars and their Spectra: An Introduction to the Spectral Sequence, Springer
27. Karoff, C., Arentoft, T., Glowienka, L., Countures, C., Nielsen, T.B., Dogan, G., Grundahl, F., Kjeldsen, H., 2008, MNRAS 386, 1085
28. Kiriakidis, M., El Eid, M.F., Glatzel, W., 1992, MNRAS 255, 1
29. Koen, C., 1993, MNRAS 264, 165
30. Kolaczowski, Z., Pigulski, A., Soszynski, I., Udalski, A., Szymanski, M., Kubiak, M., Zebrun, K., Pietrzynski, G., Wozniak, P.R., Szewczyk, O., Wyrzkowski, L., 2004, ASPC, 310, 225
31. Kolaczowski, Z., Pigulski, A., Soszynski, I., Udalski, A., Kubiak, M., Szymanski, M., Zebrun, K., Pietrzynski, G., Wozniak, P.R., Szewczyk, O., Wyrzkowski, L., 2006, Mem. S.A.It. 77, 336
32. Lesh, J.R., Aizenman, M.L., 1978, ARA&A 16, 215
33. Lynds, C.R., 1959, ApJ 129, 674
34. Maeder, A., Grebel, E.K., Mermilliod, J.-C., 1999, A&A 346, 456

35. Martayan, C., Diago, P., Gutierrez-Soto, J., Fabregat, J., Hubert, A.-M., Floquet, M., Neiner, C., Mekkas, M., 2008, IAU symposium no. 256
36. McNamara, D.H, Struve, O., Bertiau, F.C., 1955, ApJ 121, 326
37. Miglio, A., Montalban, J., Dupret, M.-A., 2007, Comm. in Asteroseismology, Vol. 151
38. Moskalik, P., Dziembowski, W.A., 1992, A&A 256, L5-8
39. Neiner, C., Hubert, A.-M., Comm. in Asteroseismology, Vol. 158
40. Osaki, Y., 1993, ASPC 40, 5120
41. Percy, J.R., 1967, JRASC 61, 117
42. Pesnell, W.D., Cox, A.N., 1980, SSRv 27, 337
43. Pigulski, A., Kolaczowski, Z., 2002, A&A 388, 88
44. Pigulski, A., 2004, Comm. Asteroseism, 145, 74
45. Pigulski, A., Pojmański, G., 2008, A&A 477, 917
46. Rogers, F.J., Iglesias, C.A., 1992, ApJS 79, 507
47. Saesen, S., Pigulski, A., Carrier, F., et al., 2010, Astron. Nachr., 331, 1080
48. Schechter, P.L., Mateo, M., Saha, A., 1993, PASP 105, 1342
49. Seaton, M.J., 1996, MNRAS 279, 95
50. Stankov, A., Handler, G., 2005, ApJS 158, 193
51. Stateva, I., Niemczura, E., Iliev, I., 2010, Publ. Astron. Obs. Belgrade No. 90, 179
52. Sterken, C., Jerzykiewicz, M., 1993, SSRv 62, 95
53. Struve, O., 1954, ApJ 119, 168

- 54. Struve, O., 1955, PASP 67, 135
- 55. Sturoock, P., Scargle, J.D., ApJ 718, 527
- 56. Torres, G., 2010, ApJ 140, 1158
- 57. Walker, M.F., 1952, AJ 57, 227
- 58. Zima, W., 2008, Communications in Asteroseismology 155

Appendix A

This part of the thesis contains all the additional information and work done during the preparation of my thesis. It consists of computer programs written to do various aspects.

A.1 Target selection

Below is the program used to select the targets with the smallest FAP from the file called olev. It is written in python programming language.

```
#####  
  
import os  
  
file = open('olev','r')  
  
file_out = open('Freq-amp.file','w')  
  
LN = 0  
  
counter = 0  
  
  
for line in file:  
    if LN >= 1:  
        ## Reading the olev file for required parameters  
        FILE = str(line.split()[0])  
        Freq = float(line.split()[4])  
        Amp = float(line.split()[5])
```

```

## Selecting required targets using Freq cut-off os 3c/d
if Freq >= 3.0:
    FAP = float(line.split()[6])

    ## further targets selection using FAP sut-off of 0.05
    if FAP <= 0.05:
        file_out.write('%s %f %f %f\n' % (FILE,Freq,Amp,FAP))
    counter += 1
    LN += 1
file.close()
file_out.close()
#####

```

The output of the above program is the file called `Freq-Amp.file` and it has four columns. The first column contains the names of the stars selected, the second column are the frequencies, the third column contains the amplitude and the fourth column contains the values of FAP.

A.2 Analysis and Results

This section gives additional work done on the chapter of analysis and the results, it starts with the description of how the stars from DoPhot output were matched to that from ISIS output in order to plot the light curves of exactly the same star from both programs.

A.2.1 Comparison of DoPhot and ISIS light curves

For comparing the light curves from DoPhot reductions to those from ISIS reductions, the reference images used on both programs had to be aligned. This was done by calculating the coordinates transformation between the reference images. For DoPhot, the image used as the reference is the one selected when `ccdcatlog` program was executed. For ISIS reductions, the reference image is produced after executing the program: `ref.csh`, and it's name is `ref.fits`. Coordinates transformation was done by solving the linear equations of the form:

$$XD = a_0 + a_1XI + a_2YI \quad (\text{A.1})$$

and

$$YD = b_0 + b_1XI + b_2YI \quad (\text{A.2})$$

as shown below. I will show the solutions of the first equation (i.e. a_0 , a_1 and a_2), because the second equation is solved in the same way. Equation to be solved is the following:

$$XD_i = a_0 + a_1XI_i + a_2YI_i \quad (\text{A.3})$$

where XD is the x-coordinate of the star on DoPhot reference image, XI and YI are x- and y-coordinates of the star on `ref.fits`. This equation is solved by minimizing the sum of the squares of the error and we are defining XD as having error (ε_i), such that the above equation can be written as:

$$XD_i + \varepsilon_i = a_0 + a_1XI_i + a_2YI_i \quad (\text{A.4})$$

Therefore, a_0 , a_1 and a_2 are solved by minimizing the sum of ε^2 as follows:

$$\varepsilon_i = a_0 + a_1XI_i + a_2YI_i - XD_i \quad (\text{A.5})$$

$$\varepsilon_i^2 = (a_0 + a_1XI_i + a_2YI_i - XD_i)^2 \quad (\text{A.6})$$

$$\sum_{i=1}^n \varepsilon_i^2 = \sum_{i=1}^n (a_0 + a_1XI_i + a_2YI_i - XD_i)^2 \quad (\text{A.7})$$

To simplify this, let E be $\sum_{i=1}^n \varepsilon_i^2$, then

$$E = \sum_{i=1}^n (a_0 + a_1XI_i + a_2YI_i - XD_i)^2 \quad (\text{A.8})$$

minimizing E with respect to a_0 , a_1 and a_2 gives:

$$\frac{\partial E}{\partial a_0} = 2 \sum_{i=1}^n (a_0 + a_1XI_i + a_2YI_i - XD_i) \quad (\text{A.9})$$

$$\frac{\partial E}{\partial a_1} = 2 \sum_{i=1}^n (a_0 + a_1XI_i + a_2YI_i - XD_i)XI_i \quad (\text{A.10})$$

and

$$\frac{\partial E}{\partial a_2} = 2 \sum_{i=1}^n (a_0 + a_1XI_i + a_2YI_i - XD_i)YI_i \quad (\text{A.11})$$

For E to be minimum, $\frac{\partial E}{\partial a_0}$, $\frac{\partial E}{\partial a_1}$ and $\frac{\partial E}{\partial a_2}$ should be equal to zeros, i.e.

$$2 \sum_{i=1}^n (a_0 + a_1XI_i + a_2YI_i - XD_i) = 0 \quad (\text{A.12})$$

$$2 \sum_{i=1}^n (a_0 + a_1XI_i + a_2YI_i - XD_i)XI_i = 0 \quad (\text{A.13})$$

$$2 \sum_{i=1}^n (a_0 + a_1XI_i + a_2YI_i - XD_i)YI_i = 0 \quad (\text{A.14})$$

Simplifying above equations gives

$$\sum a_0 + \sum a_1XI_i + \sum a_2YI_i = \sum XD_i \quad (\text{A.15})$$

$$\sum a_0 X I_i + \sum a_1 X I_i^2 + \sum a_2 Y I_i X I_i = \sum X D_i Y I_i \quad (\text{A.16})$$

$$\sum a_0 Y I_i + \sum a_1 X I_i Y I_i + \sum a_2 Y I_i^2 = \sum X D_i Y I_i \quad (\text{A.17})$$

Taking the coefficients a_0 , a_1 and a_2 out of the summation sign gives

$$a_0 + a_1 \langle X I \rangle + a_2 \langle Y I \rangle = \langle X D \rangle \quad (\text{A.18})$$

$$a_0 \langle X I \rangle + a_1 \langle X I^2 \rangle + a_2 \langle X I Y I \rangle = \langle X D X I \rangle \quad (\text{A.19})$$

$$a_0 \langle Y I \rangle + a_1 \langle X I Y I \rangle + a_2 \langle Y I^2 \rangle = \langle X D Y D \rangle \quad (\text{A.20})$$

Above equations can be written in matrix form:

$$Aa = M \quad (\text{A.21})$$

where A is 3×3 matrix of XI 's and YI 's, a is 3×1 matrix of coefficients (a_0 , a_1 , a_2) and M is 3×1 matrix of XD 's and YD 's. The coefficients can be found by finding the solutions of matrix multiplication:

$$a = (A^{-1})M \quad (\text{A.22})$$

The solutions of the above matrix were obtained with python program which is attached below.


```
#####
from numpy import *
print 'This program finds the coefficients on the equation of the form:'
print ' y = a0 + a1*x1 + a2*x2 + ... + am*xm  '
print 'Input data file should be in form x1, x2, ..., xm, y'
indv = int(raw_input("Number of independent variables? "))
indv1 = indv + 1
fp = open(raw_input("Enter the file containing coordinates: "), "r")
x = zeros((indv1, indv1))
y = zeros((indv1, 1))
xi = [0.0]*indv1
n = 0
for line in fp:
    xi[0] = 1.0
    for j in range(1, indv1):
        xi[j] = float(line.split()[j-1])
    yi = float(line.split()[indv])
    for j in range(indv1):
        for k in range(indv1):
            x[j, k] += xi[j]*xi[k]
        y[j, 0] += yi*xi[j]
    n += 1
fp.close()
X = mat( x.copy() )
Y = mat( y.copy() )
A = X.I*Y
for j in range(indv1):
    print "A[%d] = %f" % (j, A[j])
#####
```

In order for this program to work properly, the input file should be having columns in the following order:

XI YI XD

and there are two independent variables (i.e XI and YI). The entries of the input file are the (x,y) coordinates of the stars on the two reference frames mentioned above. It should be noted that XI_i and YI_i on `ref.fits` must be for the same star with coordinates XD_i and YD_i on the reference image used on Dophot reductions. This means that the star with XI_i, YI_i on `ref.fits` must be the same star with XD_i, YD_i on DoPhot's reference image.

The python program requires the input data file with three columns as shown above, and the linear equations we are solving imply that one needs to find the coordinates transformation of DoPhot's reference frame relative to ISIS's reference frame. This means that after getting the values of coefficients, we use them to find transformed coordinates of stars on `ref.fits` and compare them with the coordinates on `catalog.dat` file. In order to get all values of coefficients, one should first include column three of input file as XD_i (this will give (a_0 , a_1 and a_2)), and after getting results, edit the third column to YD_i (which will give values of (b_0 , b_1 and b_2)).

The transformed coordinates were used to match ISIS stars to DoPhot stars to two pixels. The program used to do coordinates matching is shown below.

```
#####
file1 = open(raw_input('Enter file of ISIS stars: '), 'r')
file2i = str(raw_input('Enter file of DOPHOT stars: '))
file3 = open(raw_input('Enter output file: '), 'w')
file3.write("%s\n%s\n%s\n%s\n%s\n%s\n%s\n%s\n" % ("ISIS file", "c1-X ISIS"
        , "c2-Y ISI", "c3-trans-X", "c4-trans-Y", "c5-DOPHOT file#", "c6-X DoP", "c6-
        Y DoP"))

mat = int(raw_input('Enter pixels number for matching coordinates: '))
a0 = float(raw_input("enter a0: "))
a1 = float(raw_input("enter a1: "))
a2 = float(raw_input("enter a2: "))
b0 = float(raw_input("enter b0: "))
b1 = float(raw_input("enter b1: "))
b2 = float(raw_input("enter b2: "))

for line in file1:
    X1 = float(line.split()[0])
    Y1 = float(line.split()[1])
    LC = str(line.split()[4])

## calculating the transformed X,Y coordinates of X1 and Y1
    tX1 = a0 + a1*X1 + a2*Y1
    tY1 = b0 + b1*X1 + b2*Y1

## comparing the transformed X1,Y1 to DoPhot's output
    file2 = open(file2i, 'r')

    for line in file2:
        X2 = float(line.split()[2])
        Y2 = float(line.split()[3])
        STR_no = int(line.split()[1])
```


C

```
if (tX1-mat <= X2 <= tX1+mat) and (tY1-mat <= Y2 <= tY1+mat):
```

```
    file3.write("%s %f %f %f %f %d %0.2f %0.2f\n"
```

```
                % (LC,X1,Y1,tX1,tY1,STR_no,X2,Y2))
```

```
file1.close()
```

```
file2.close()
```

```
file3.close()
```

```
#####
```

The output file after the above program was executed has the columns ordered as follows:

c1-ISIS file
c2-X ISI
c3-Y ISI
c4-trans-X
c5-trans-Y
c6-DOPHOT file#
c7-X DoP
c8-Y DoP

```
=====
lc0.data  14.557054  211.846503  91.844501  132.839270    5523  91.53   130.81
lc1.data  17.325245  452.948403  94.561573  373.651095     8    96.16   375.37
lc2.data  22.785881  81.990793 100.128776  3.160854    1885 102.45    2.49
lc3.data  24.920740  483.801630 102.171323 404.482894    44   102.63  402.19
lc4.data  25.781230  479.500835 103.035338 400.189245    44   102.63  402.19
lc4.data  25.781230  479.500835 103.035338 400.189245    343  104.74  398.55
:         :         :         :         :         :         :
:         :         :         :         :         :         :
```

where columns 1-8 are ISIS file name, (x,y) coordinates from ISIS reference frame, transformed (x,y) coordinates, DoPhot star/file number and (x,y) coordinates from DoPhot reference frame.

A.2.2 “cat” files manipulations

Since the correction of the positions of the stars in `sum` files should be done with all the files of all filters together, we had to write a program which will separate the resulting `cat` files. Such program is shown below and it works by editing the third `if` statement with an appropriate filter number. The `cat` files for other filters are being deleted as the fourth to sixth `if` statements show. This program was used by creating four directories for U,B,V and I filters and copying all the produced `cat` files to each of the directories. Then the program was executed in each of the directories.

```
#####  
import os  
CAT = int(raw_input('Enter total number os sum files: '))  
for i in range(CAT):  
    j = 10000 + i  
    catfile = 'cat.00' + str(j)  
    k = 0  
    if os.path.exists(catfile):  
        file1 = open(catfile,'r')  
        for line in file1:  
            k = k + 1  
            if k == 3:  
                filt = int(line.split()[1])  
                if filt == 18:  
                    continue  
                elif filt == 38:  
                    os.remove(catfile)  
                elif filt == 28:  
                    os.remove(catfile)  
                elif filt == 58:  
                    os.remove(catfile)
```



```
file1.close()
```

```
#####
```

A.2.3 Standardization of magnitudes

A python program below was used to calculate the distance ratios on the CCD image and on OGLE II maps corresponding to our images. This program takes an input file which contains the x,y coordinates of the stars in CCD (columns 2 and 3) and corresponding stars in OGLE II maps (columns 4 and 5). The distance ratios are compared to confirm that the stars identified are corresponding stars in both CCD and OGLE II images.

```
#####
```

```
from math import *
```

```
def distance(x1,y1,x2,y2):
```

```
    A = (x2-x1)**2
```

```
    B = (y2-y1)**2
```

```
    D = sqrt(A+B)
```

```
    return D
```

```
#####
```

```
f1 = open('coord.in','r')
```

```
strn = []
```

```
x0 = []
```

```
y0 = []
```

```
xD = []
```

```
yD = []
```

```

for line in f1:
    strn.append(int(line.split()[0]))
    x0.append(float(line.split()[3]))
    y0.append(float(line.split()[4]))
    xD.append(float(line.split()[1]))
    yD.append(float(line.split()[2]))

Ogle_D = []
ccd_D = []

for i in range(1,len(strn)):
    d1 = distance(xD[0],yD[0],xD[i],yD[i])
    d2 = distance(x0[0],y0[0],x0[i],y0[i])
    ccd_D.append(d1)
    Ogle_D.append(d2)

print 'CCD distances = ',ccd_D
print 'Ogle II distances = ', Ogle_D

for j in range(len(ccd_D)):
    print ccd_D[0]/ccd_D[j], Ogle_D[0]/Ogle_D[j]

f1.close()

#####

```

A python program below was used to obtain RA and DEC coordinates of the stars on the template images. This was done in order to match stars from template images to those on same field of the OGLE, such that the magnitudes of the OGLE stars can be used to standardize the magnitudes of stars on the template images. The matching was first done by finding few stars which can be recognized from both the OGLE fields and our template images. The coordinates of stars found were written in some file in

the form:

$$X_i \ Y_i \ RA_i \ DEC_i$$

where X_i and Y_i are template image coordinates of the stars, RA_i and DEC_i are the coordinates from the OGLE fields images. The file containing these coordinates is an input to a Fortran program which was provided by Dr. Balona to align stars on template images to those on the OGLE fields images. Produced alignment parameters are used by python program below together with two files containing stellar parameters from both our reductions and the OGLE. The python program uses alignment parameters to match stars on two files and write out the magnitudes and coordinates of matched stars in one output file. The stars in this output file have magnitude values from our reductions and from the OGLE database. These magnitudes are used for standardization as explained in chapter 4.


```
#####
```

```
from math import*
Up = 12000*[0.0]
Bp = 12000*[0.0]
Vp = 12000*[0.0]
Ip = 12000*[0.0]
fp = open('phot','r')
for line in fp:
    id = int(line.split()[0])
    Up[id] = float(line.split()[1])
    Bp[id] = float(line.split()[2])
    Vp[id] = float(line.split()[3])
    Ip[id] = float(line.split()[4])
fp.close()
file1 = open('radec.out','r')
file2 = open('ogle.in','r')
fo2 = open('match.out','w')
RA = []
DEC = []
strn1 = []
num = 0
dmax = -1.0e+20
dmin = 1.0e+20
rmax = -1.0e+20
rmin = 1.0e+20
for line in file1:
    strn1.append(int(line.split()[0]))
    ra1 = str(line.split()[5])
    hr = float(ra1.rsplit(':')[0])
    min = float(ra1.rsplit(':')[1])
```

```

sec = float(ra1.rsplit(':')[2])
r = hr + min/60.0 + sec/3600.0
RA.append(r)
if r > rmax:
    rmax = r
if r < rmin:
    rmin = r
dec1 = str(line.split()[6])
deg = float(dec1.rsplit(':')[0])
min = float(dec1.rsplit(':')[1])
sec = float(dec1.rsplit(':')[2])
d = abs(deg) + min/60.0 + sec/3600.0
if deg < 0.0:
    d = -d
DEC.append(d)
if d > dmax:
    dmax = d
if d < dmin:
    dmin = d
num += 1
print rmin,rmax,dmin,dmax
tol = 1.0
for line in file2:
    strn2 = int(line.split()[1])
    ra2 = float(line.split()[2])

    if (ra2 >rmin) and (ra2<rmax) :
        dec2 = float(line.split()[3])
        if (dec2>dmin) and (dec2<dmax) :
            V = float(line.split()[6])

```

```

I = float(line.split()[8])
B = float(line.split()[10])
dsmin = 1.0e+20
fac = 15.0*cos(radians(dec2))
for i in range(num):
    dra = fac*(ra2 - RA[i])
    ddec = dec2 - DEC[i]
    ds = dra*dra + ddec*ddec
    if ds < dsmin:
        dsmin = ds
        imin = i
ds = 3600.0*sqrt(dsmin)
if ds < tol :
    id = strn1[imin]
    fo2.write('%6d %6.1f %8.3f %8.3f %8.3f %8.3f %6d %8.3f
%8.3f %8.3f\n' % (id,ds,Up[id],Bp[id],Vp[id],Ip[id],strn2,B,V,I))
file1.close()
file2.close()
fo2.close()
#####

```

The last part of standardization of our magnitudes was done with a series of python-written programs. These programs perform analysis to eventually producing HR-diagrams for the stars in two fields we observed. Each of these analysis programs is attached with the description below.

Program A: The following program reads the file produced by matching stars in our frames to those in corresponding OGLE II frames. It saves star numbers from our reductions, magnitudes from both our reductions and OGLE II and magnitude differences in an output file for each of BVI filters.

```
#####
```



```

from math import *

# Reading the data file.

f1 = open(raw_input('Enter input filename: '), 'r')
f2b = open('b.out', 'w')
f2v = open('v.out', 'w')
f2i = open('i.out', 'w')

B = []
b = []
V = []
v = []
I = []
ii = []
strn = []

for line in f1:
    strn.append(int(line.split()[0]))
    b.append(float(line.split()[3]))
    B.append(float(line.split()[7]))
    v.append(float(line.split()[4]))
    V.append(float(line.split()[8]))
    ii.append(float(line.split()[5]))
    I.append(float(line.split()[9]))

for i in range(len(ii)):
    f2b.write('%d %8.4f %8.4f %8.4f\n' % (strn[i], b[i], B[i], B[i]-b[i]))
    f2v.write('%d %8.4f %8.4f %8.4f\n' % (strn[i], v[i], V[i], V[i]-v[i]))
    f2i.write('%d %8.4f %8.4f %8.4f\n' % (strn[i], ii[i], I[i], I[i]-ii[i]))

f1.close()
f2b.close()
f2v.close()
f2i.close()

#####

```

Program B: The program below reads the output files from the program A above. Before program A outputs can be used, the stars in those files should be sorted in order of their decreasing brightness, with brightest stars on the first rows. This is simply done with a statement:

```
sort -g -k [column no.] sorted file > output file
```

This statement is executed on the command line and in our analysis we sorted stars according to OGLE II magnitudes. The program below reads sorted files, calculates the average values and standard deviation values of the specified number of brightest stars, and print out calculated values. The output file contains the star numbers and magnitude differences in B, V and I filters.

```
#####  
from math import *  
from numpy import *  
# Function to calculate Standard Deviation.  
def stddev(A):  
    Xx = 0.0  
    n = 0  
    for i in range(len(A)):  
        Xx = Xx + (A[i] - mean(A))**2  
        n = n + 1  
    std = sqrt(Xx/(n - 1))  
    return std  
# Opening input files  
fb = open('b.dat','r')  
fv = open('v.dat','r')  
fi = open('i.dat','r')  
delB = []  
strn = []  
delV = []
```

```

delI = []

for line in fb:
    strn.append(int(line.split()[0]))
    delB.append(float(line.split()[3]))

for line in fv:
    strn.append(int(line.split()[0]))
    delV.append(float(line.split()[3]))

for line in fi:
    strn.append(int(line.split()[0]))
    delI.append(float(line.split()[3]))

DELb = []
DELv = []
DELi = []

ffout = open('BS_magdiff.dat','w')

stars = int(raw_input('Enter number of brightest stars to use: '))

for i in range(stars):
    DELb.append(delB[i])
    DELv.append(delV[i])
    DELi.append(delI[i])
    ffout.write('%d %8.4f %8.4f %8.4f\n'
                % (strn[i],delB[i],delV[i],delI[i]))

ffout.close()

stdV = stddev(DELv)
avgV = sum(DELv)/float(stars)
stdI = stddev(DELi)
avgI = sum(DELi)/float(stars)
stdB = stddev(DELb)
avgB = sum(DELb)/float(stars)

print 'stdB = %8.4f' % (stdB), 'and avgB = %8.4f' % (avgB)
print 'stdV = %8.4f' % (stdV), 'and avgV = %8.4f' % (avgV)

```



```
print 'stdI = %8.4f' % (stdI), 'and avgI = %8.4f' % (avgI)
```

```
#####
```

Program C: After brightest stars were written in some output file, the program below reads that file and also intakes calculated average magnitude difference values and standard deviations. It ignores the stars with a difference (between magnitude difference and average values) exceeding three times standard deviation. This is done for all B, V and I magnitudes. The program then saves the remaining stars in some output files.

```
#####
```

```
from math import *

f1 = open(raw_input('Enter brightest stars file: '), 'r')
f2b = open('B1.dat', 'w')
f2v = open('V1.dat', 'w')
f2i = open('I1.dat', 'w')

stdB = float(raw_input('Enter std. dev. of B: '))
stdV = float(raw_input('Enter std. dev. of V: '))
stdI = float(raw_input('Enter std. dev. of I: '))
avgB = float(raw_input('Enter B-average difference: '))
avgV = float(raw_input('Enter V-average difference: '))
avgI = float(raw_input('Enter I-average difference: '))

for line in f1:
    strn = int(line.split()[0])
    Bdiff = float(line.split()[1])
    Vdiff = float(line.split()[2])
    Idiff = float(line.split()[3])
    if abs(Bdiff - avgB) < 3*stdB:
        f2b.write('%d %8.4f %8.4f %8.4f\n'
                  % (strn, Bdiff, Bdiff-avgB, 3*stdB))
    if abs(Vdiff - avgV) < 3*stdV:
```

```

f2v.write('%d %8.4f %8.4f %8.4f\n'
          % (strn,Vdiff,Vdiff - avgV,3*stdV))
if abs(Idiff - avgI) < 3*stdI:
    f2i.write('%d %8.4f %8.4f %8.4f\n'
              % (strn,Idiff,Idiff - avgI,3*stdI))

f1.close()
f2b.close()
f2v.close()
f2i.close()

#####

```

Program D: The following program does what the program B does, except that it intakes the outputs from the program C. This allows one to manually removes any stars which deviates most from the average.

```

#####

from math import *
from numpy import *

# Function to calculate Standard Deviation.
def stddev(A):
    Xx = 0.0
    n = 0
    for i in range(len(A)):
        Xx = Xx + (A[i] - mean(A))**2
        n = n + 1
    std = sqrt(Xx/(n - 1))
    return std

f2b = open('B1.dat','r')
fb = open('B2.dat','w')
f2v = open('V1.dat','r')
fv = open('V2.dat','w')

```

```

f2i = open('I1.dat','r')
fi = open('I2.dat','w')
Bdiff = []
strn = []
Vdiff = []
Idiff = []
nB = 0.0
nV = 0.0
nI = 0.0
for line in f2b:
    strn.append(int(line.split()[0]))
    Bdiff.append(float(line.split()[1]))
    nB += 1
for line in f2v:
    strn.append(int(line.split()[0]))
    Vdiff.append(float(line.split()[1]))
    nV += 1
for line in f2i:
    strn.append(int(line.split()[0]))
    Idiff.append(float(line.split()[1]))
    nI += 1
avgB = sum(Bdiff)/nB
avgV = sum(Vdiff)/nV
avgI = sum(Idiff)/nI
stdB = stddev(Bdiff)
stdV = stddev(Vdiff)
stdI = stddev(Idiff)
print 'stdB = %8.4f' % (stdB), 'and avgB = %8.4f' % (avgB)
print 'stdV = %8.4f' % (stdV), 'and avgV = %8.4f' % (avgV)
print 'stdI = %8.4f' % (stdI), 'and avgI = %8.4f' % (avgI)

```



```

for i in range(len(Bdiff)):
    if abs(Bdiff[i] - avgB) < 3*stdB:
        fb.write('%d %8.4f %8.4f %8.4f\n'
                  % (strn[i],Bdiff[i],Bdiff[i]-avgB,3*stdB))
for i in range(len(Vdiff)):
    if abs(Vdiff[i] - avgV) < 3*stdV:
        fv.write('%d %8.4f %8.4f %8.4f\n'
                  % (strn[i],Vdiff[i],Vdiff[i]-avgV,3*stdV))
for i in range(len(Idiff)):
    if abs(Idiff[i] - avgI) < 3*stdI:
        fi.write('%d %8.4f %8.4f %8.4f\n'
                  % (strn[i],Idiff[i],Idiff[i]-avgI,3*stdI))

fb.close()
f2b.close()
f2v.close()
fv.close()
f2i.close()
fi.close()

#####

```

Program E: The program below accept as input an average values of magnitude differences calculated from the brightest stars with smallest deviations. It uses these values to correct all the B, V and I magnitudes from our reductions.

```

#####
fin = open(raw_input('Enter file containing magnitudes to be corrected: '), 'r')
fout = open(raw_input('Enter corrected magnitudes filename for this field: '), 'w')
Vcorr = float(raw_input('Enter V-mag correction: '))
Bcorr = float(raw_input('Enter B-mag correction: '))
Icorr = float(raw_input('Enter I-mag correction: '))
for line in fin:

```

```

strn = int(line.split()[0])
True_Bmag = (float(line.split()[2])) + Bcorr
True_Vmag = (float(line.split()[3])) + Vcorr
True_Imag = (float(line.split()[4])) + Icorr
fout.write('%d %8.4f %8.4f %8.4f\n' % (strn,True_Bmag,True_Vmag,True_Imag))
fin.close()
fout.close()
#####

```

Program F: The following program uses corrected magnitudes to calculates absolute magnitudes of B, V and I apparent magnitudes. It saves the star numbers, corrected apparent magnitudes and their calculated absolute magnitudes in an output file.

```

#####
from math import *
fin = open(raw_input('enter corrected magnitude input file: '), 'r')
fout = open(raw_input('enter absolute magnitude output file: '), 'w')
d = 18.52 # distance modulus of LMC from Bonanos et al (2011)
for line in fin:
    strn = int(line.split()[0])
    Bmag = float(line.split()[1])
    Vmag = float(line.split()[2])
    Imag = float(line.split()[3])
    Mb = Bmag - d
    Mv = Vmag - d
    Mi = Imag - d
    fout.write('%d %8.4f %8.4f %8.4f %8.4f %8.4f\n'
               % (strn,Bmag,Vmag,Imag,Mb,Mv,Mi))
fin.close()
fout.close()
#####

```

The above programs from A to F could have been combined to one big program. However, stepwise analysis was found more convenient in terms of involvement understanding of how analysis is done.

A.3 Observations

This section gives the additional information about the observations related aspects. The following section gives the visibility curves of the targets we observed.

A.3.1 Visibility Curves Of Observed Targets

Due to the limitation of the 1.0-m telescope on accessing the targets at certain altitudes, it was very important to know at what time will the target to be observed be high enough for the telescope to access it. During all three observation campaigns, the 1.0-m telescope was not allowed to point the target which was below the altitude of 35° , this was for security reasons to avoid bumping the telescope against the dome structures. This precaution was fulfilled by using the visibility curves for each target we observed and these are shown below. The visibility curves of targets T1, T13 and T6 for each night these targets were observed are given, starting with the visibility curve of the first night for each target.

Visibility Curves for T1 (November 2010)

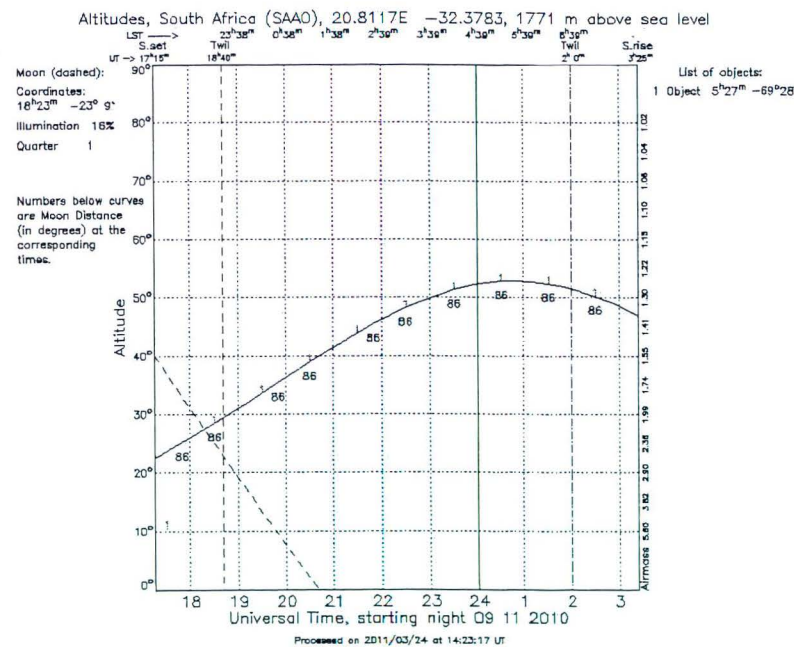
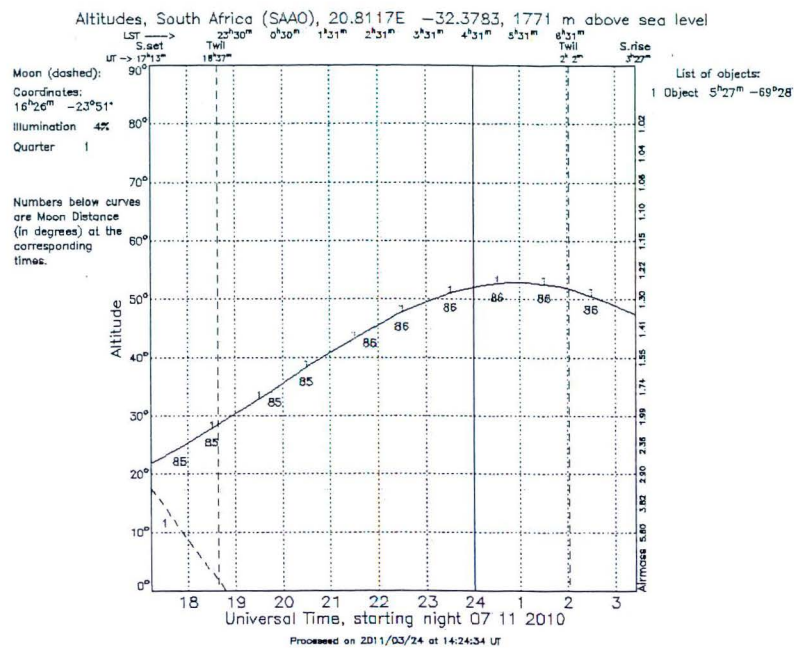


Figure A.1: Plots showing altitude as a function of time, the solid curve marks the altitude of T1 above the horizon and the dashed curve marks the altitude of the moon (taken from <http://catserver.ing.iac.es/staralt/>).

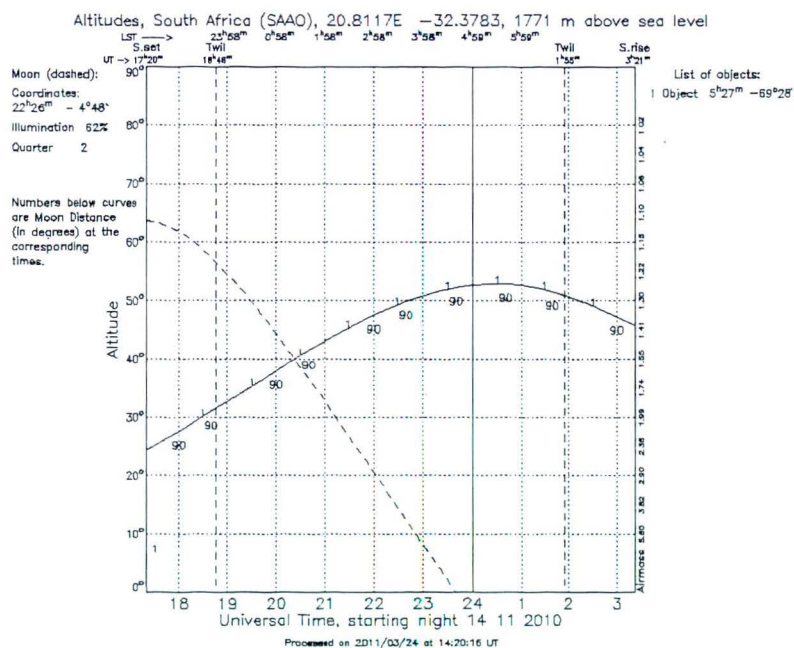
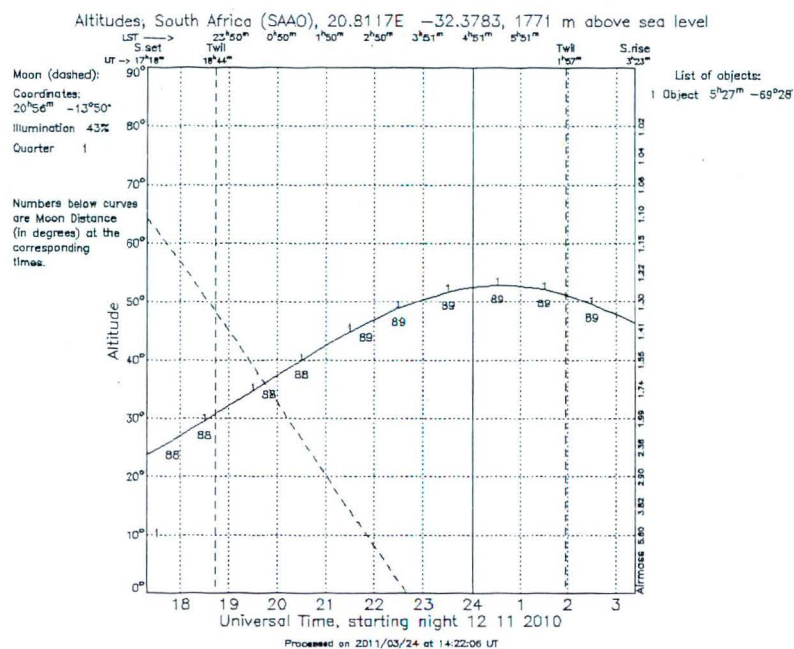


Figure A.2: Plots showing altitude as a function of time, the solid curve marks the altitude of T1 above the horizon and the dashed curve marks the altitude of the moon (taken from <http://catserver.ing.iac.es/staralt/>).

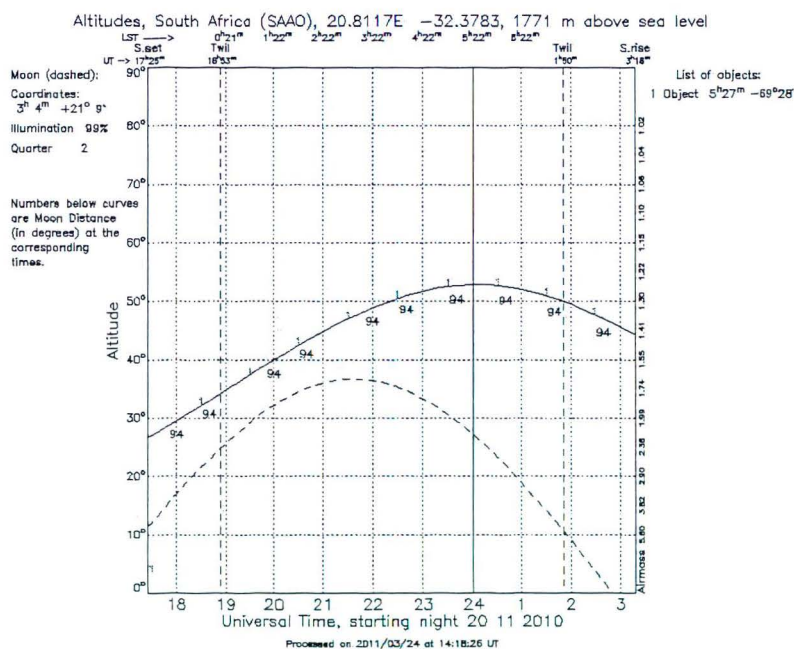
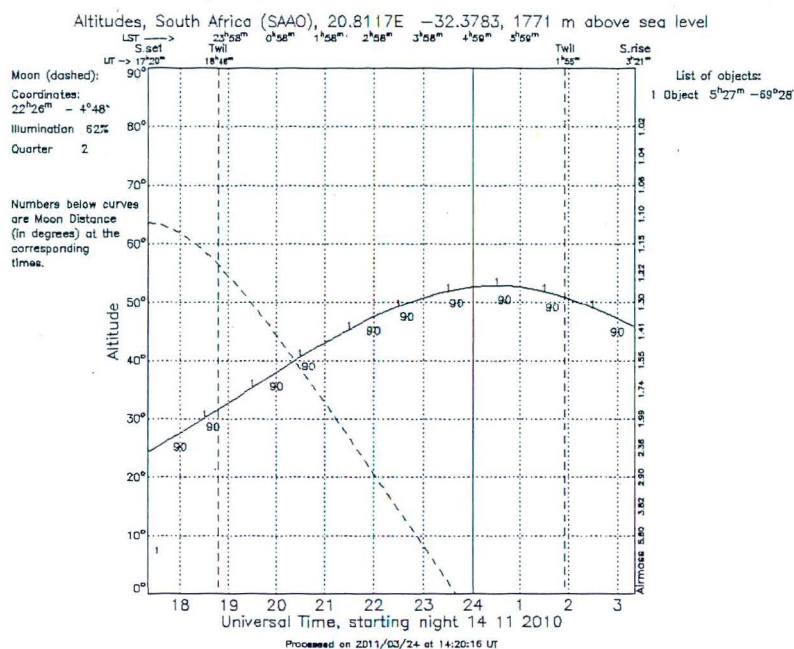


Figure A.3: Plots showing altitude as a function of time, the solid curve marks the altitude of T1 above the horizon and the dashed curve marks the altitude of the moon (taken from <http://catserver.ing.iac.es/staralt/>).

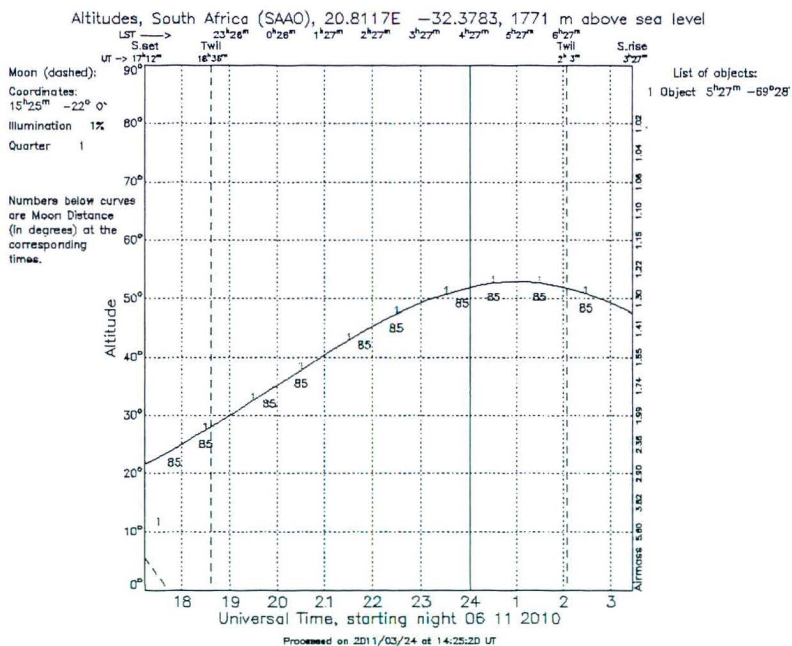
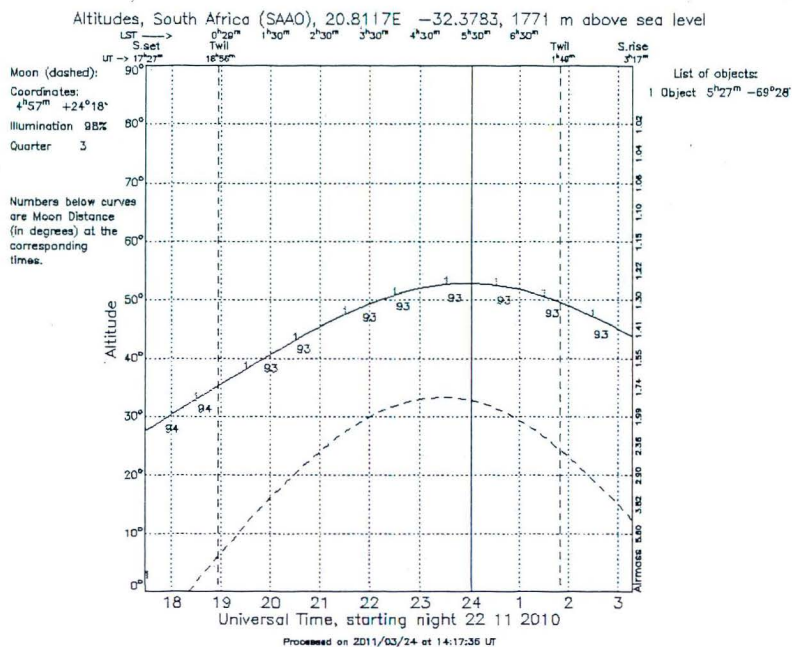


Figure A.4: Plots showing altitude as a function of time, the solid curve marks the altitude of T1 above the horizon and the dashed curve marks the altitude of the moon (taken from <http://catserver.ing.iac.es/staralt/>).

Visibility Curves for T13 (November 2010)

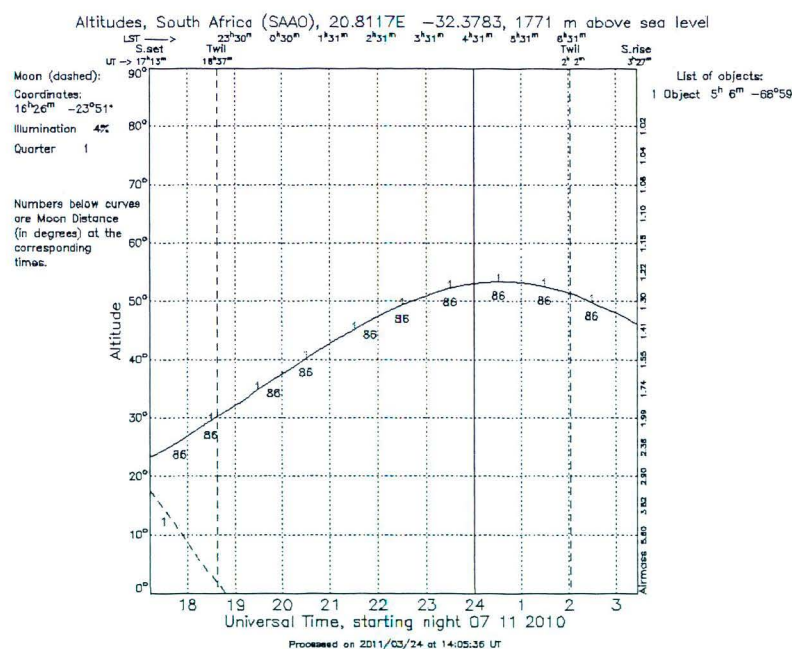
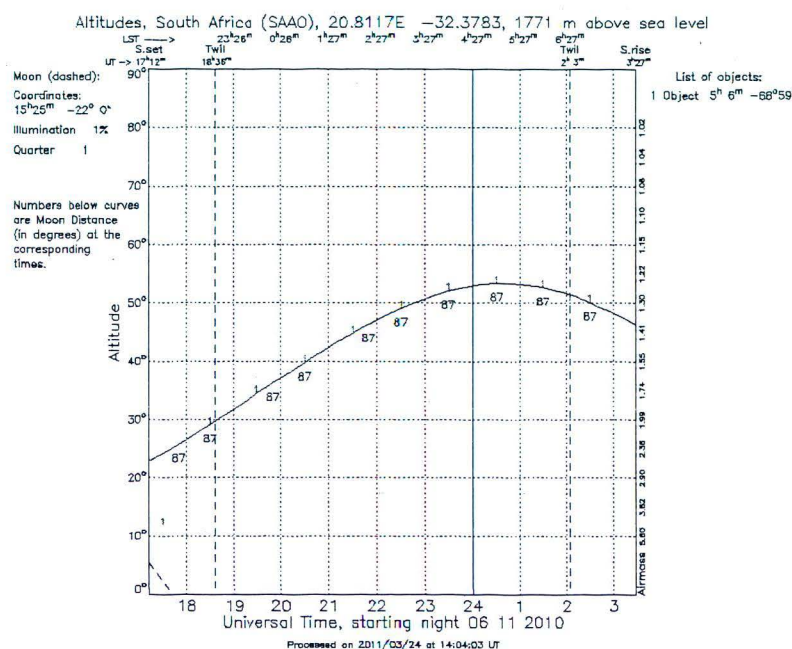


Figure A.5: Plots showing altitude as a function of time, the solid curve marks the altitude of T13 above the horizon and the dashed curve marks the altitude of the moon (taken from <http://catserver.ing.iac.es/staralt/>).

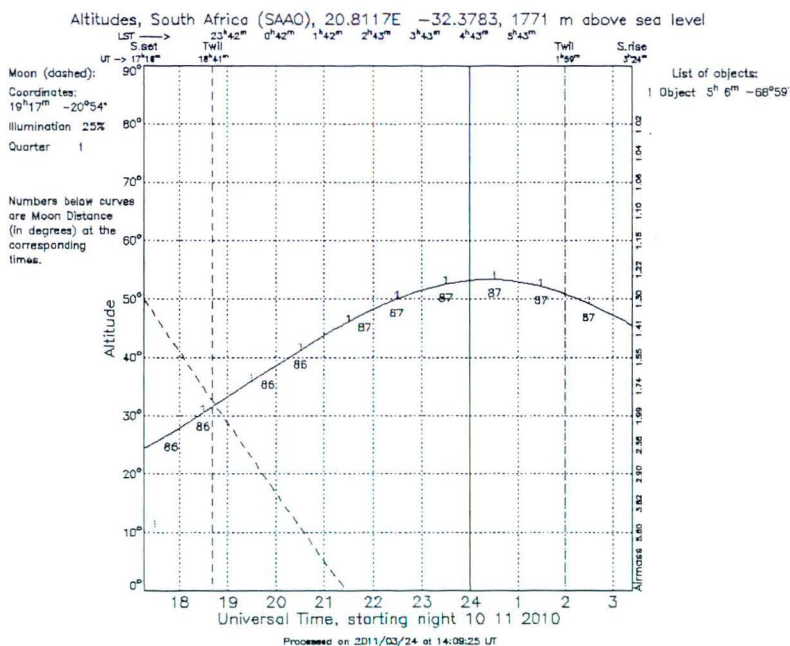
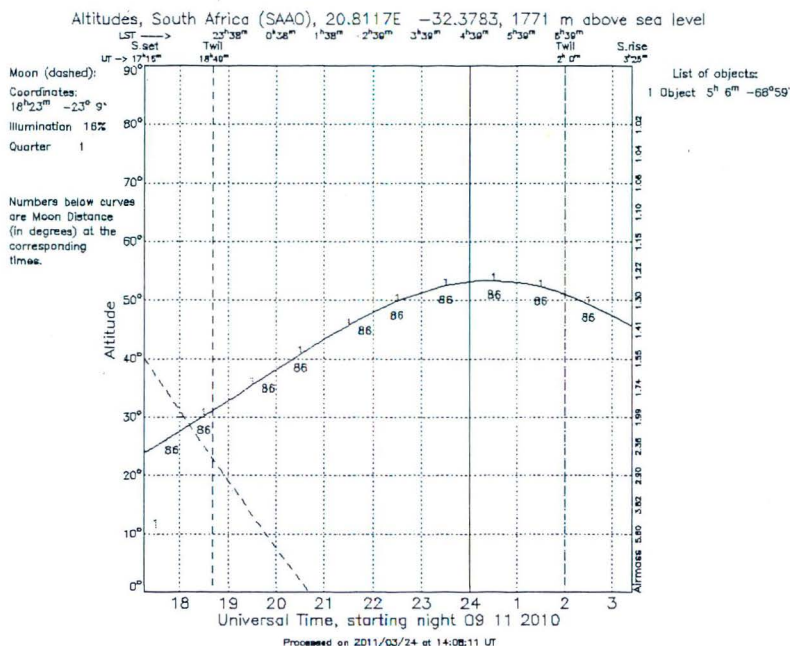


Figure A.6: Plots showing altitude as a function of time, the solid curve marks the altitude of T13 above the horizon and the dashed curve marks the altitude of the moon (taken from <http://catserver.ing.iac.es/staralt/>).

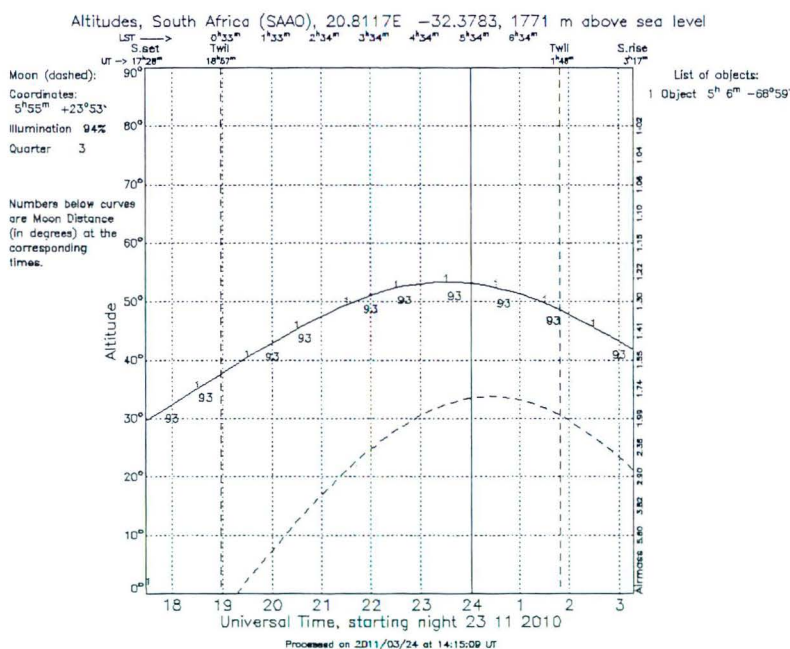
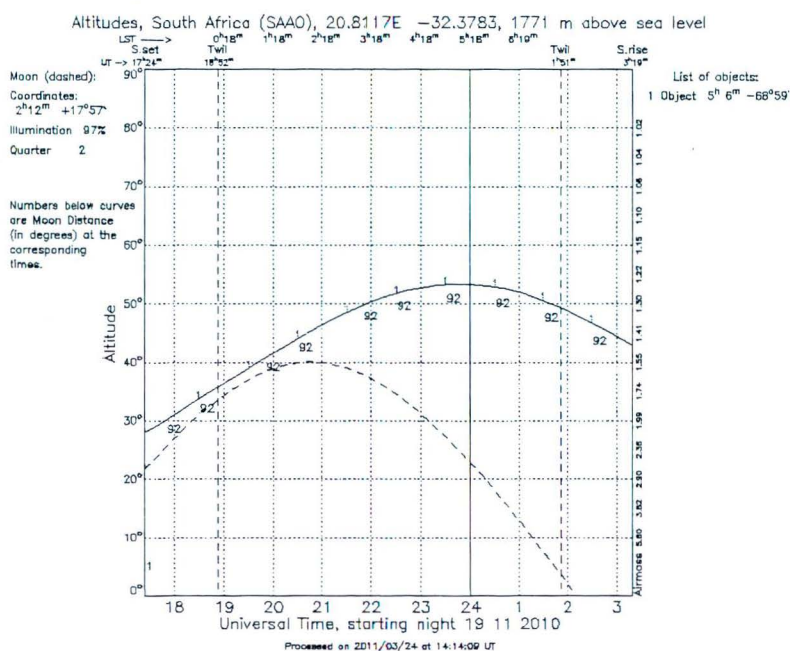


Figure A.8: Plots showing altitude as a function of time, the solid curve marks the altitude of T13 above the horizon and the dashed curve marks the altitude of the moon (taken from <http://catserver.ing.iac.es/staralt/>).

Visibility Curves for T6 (March 2011)

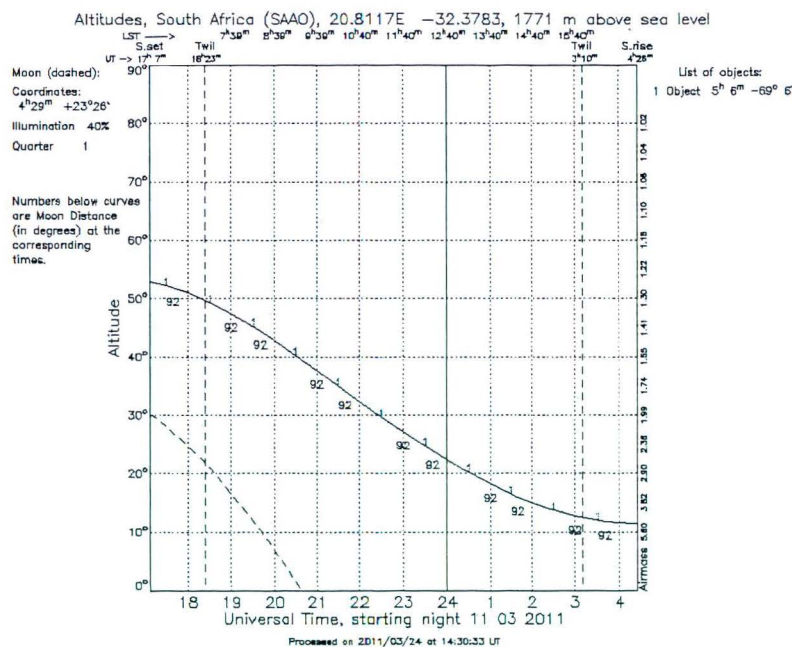
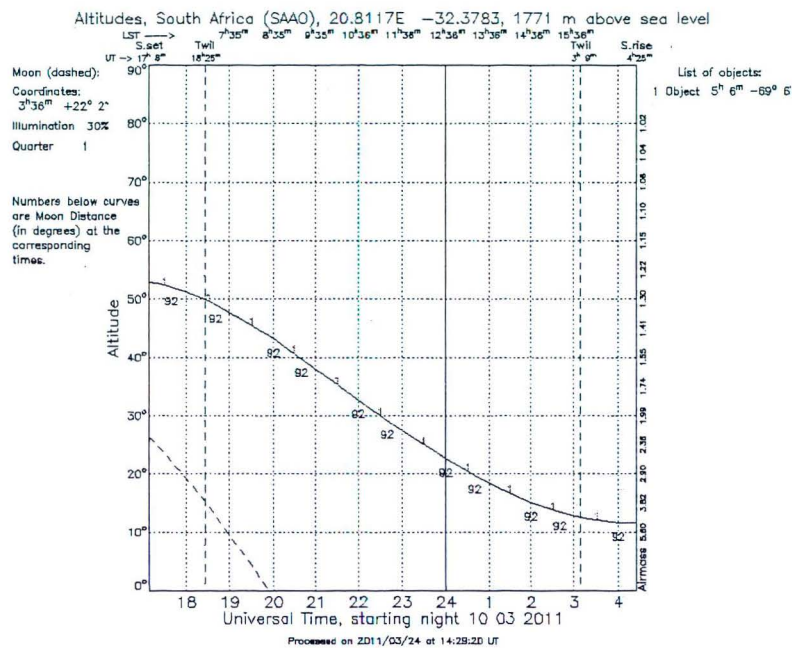


Figure A.9: Plots showing altitude as a function of time, the solid curve marks the altitude of T6 above the horizon and the dashed curve marks the altitude of the moon (taken from <http://catserver.ing.iac.es/staralt/>).

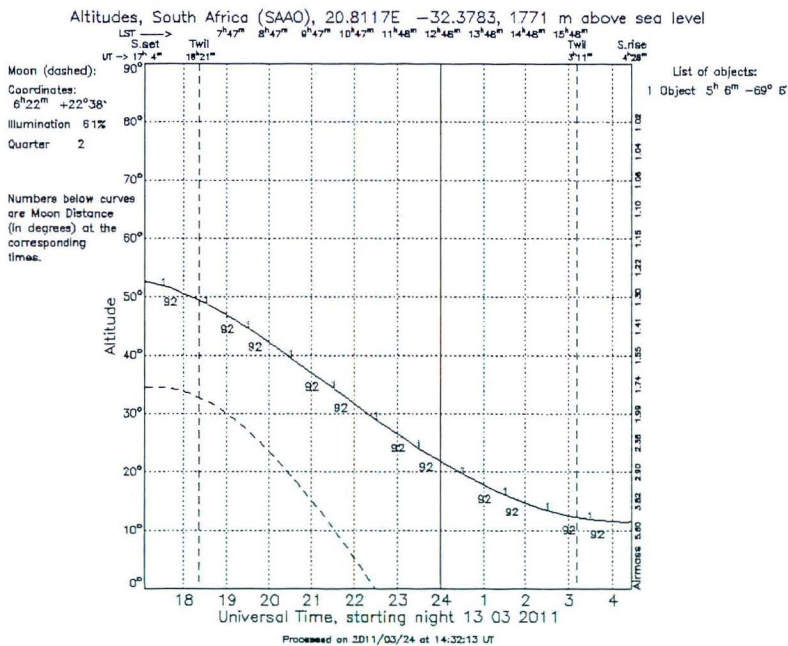
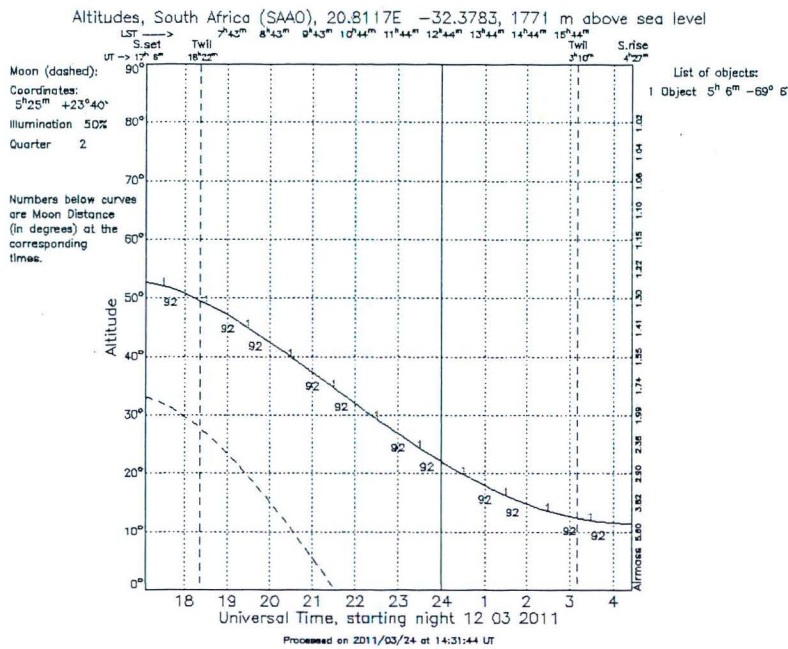


Figure A.10: Plots showing altitude as a function of time, the solid curve marks the altitude of T6 above the horizon and the dashed curve marks the altitude of the moon (taken from <http://catserver.ing.iac.es/staralt/>).

These visibility curves are the plots of Altitude against time (Universal Time [UT]), and also the airmass on the right vertical axis against the time counted from zero after Sun-set until Sun-rise on above horizontal axis. The dash vertical lines (two around 18:00 UT and 03:00 UT) mark the evening and morning twilights respectively, within this interval it is astronomically safe to proceed with observations if weather conditions are good as well. The solid curve with numbers along it shows the altitude of the target as the night progress. The dash-curve shows the altitude of the moon during that night and lastly is the solid line at 24:00 UT, which marks the mid-night. The coordinates of the target whose altitude is shown are printed on the top-right of the visibility curve and the coordinates as well as the phase of the moon are shown on top-left of the visibility curve.

A.4 Data Reductions

A.4.1 Contents of dophot.inp

Below are the contents of `dophot.inp`, which is the file containing necessary information about the CCD which was used during observations. This file is used when reducing the data using DoPhot program.

```
2  TEK8 tuneable data for DOPHOT
1  5.60          eperdn    SAAO: Electrons per digital number
2  12.32         rnoise    SAAO: Readout noise per pixel in electrons
3  -4000         ibot      SAAO: Lowest allowable data value
4  65500         itop      SAAO: Highest allowable data value
5  64000         cmax      SAAO: Central intensity threshold for obliteration
6  400.0         cmin      SAAO: 4-parameter threshold.
7  200.0         ufactor   SAAO: Normalization factor for CHISQ fit
8  3             ixby2     SAAO: Half rectangle for TRANSMASK; side = 2*ixby2+1
9  3             iyby2     SAAO: Half rectangle for TRANSMASK
10 9.0           fitfac    Conversion of Sig2x to fit raster size
11 10           nit       Maximum number iterations for fits
12 7            npar      Number of profile parameters
13 4            nfit1     Free parameters when shape predetermined
14 7            nfit2     Free parameters when fitting for shape
15 0.6          fac       Add this fraction of sub'd star to noise
16 1.6          xpnd      Treat sub'd star as bigger to compute noise
17 1.2          factor    Some margin in computing limits, from cephs
18 1.0          asprat    Compression factor along y direction
19 1            nphsub     Limiting surf brightness in dn for starsub
20 1.0          phob      Limiting surf brightness in dn for oblit
21 1            icrit     Sat'd pixel threshold for obliteration
22 0            n0right    Ignore hi pix in computing sat'd pixels
23 0            n0left     Ignore lo pix in computing sat'd pixels
```

24	16.0	chicrit	Critical chi (3 deg free) for galaxy
25	1.0	stograd	Chisquared criterion (2 stars or galaxy)
26	25.0	xtra	We ask a bit more when not all pix are present
27	0.10	sig(1)	Minimum fractional scatter in gxwid
28	10.0	sig(2)	Minimum fractional scatter in tilt
29	0.10	sig(3)	Minimum fractional scatter in gywid
30	0.50	enuff4	Min filling factor for fitting subrasters
31	.6666667	enuff7	Min filling factor for fitting subrasters
32	31	krect(1)	FIT rectangle on first pass
33	31	krect(2)	Likewise, larger to accommodate sat'd images
34	50	crit7	Lower limit for sigma**2 for 7 param fit
35	0	nobad	No of bad pixels to accept for aperture mag
36	4.0	bumpcrit	Minimum sigma thru mask to be called star
37	0.9	widobl	Side of square obliterated for cosmic ray
38	1.0	discrim	Cosmic if chi .lt. discrim * chi star
39	0	nbadleft	Eliminate low values of fast parameter
40	0	nbadright	Eliminate high values of fast parameter
41	0	nbadbot	Eliminate low values of slow parameter
42	0	nbadtop	Eliminate high values of slow parameter
43	9	sn2cos	Square of s/n to be called a cosmic ray
-1	0		END OF DATA

A.4.2 The header of the “sum” files

The output files after running dophot program have the header which contains the parameters shown below,

```
DoPHOT 5.2           Feb 98      TEK8           Tue May  3 18:26:08 2011
File   Object        Fl    RA      Dec   Date
c0010180 LMC_T13      58 05:07:05 -68:58:21 06/11/10 00:16:240   300.00
Dimension:    512 X   512    Model = 1
Sky parameters:  3    0.2102843E+04  -0.3148244E+02  -0.1185318E+02
PSF coeffs:      0.3700485E+01  -0.8543384E-03   0.3981093E+01

Sig2x   Sig2y      Sky  Fit  Apr      Tmin      Tmax      Cmin
4.65    4.25    2092.00  19  26.    134.60  15726.40  400.00
```

This header gives information on the object (RA, Dec, UT of start, exposure time etc) obtained from the FITS header. This is a frame containing 512 × 512 pixels. The “Model” is a code for how the sky is fitted. Model 1 means that the sky is assumed to be gradually changing and can be modelled as a plane. This is the normal selection and should not be changed. The Sky parameters give the coefficients of the plane representing the sky background as a function of pixel number (x,y):

$$\text{Sky} = 2102.84 - 31.48244x - 11.85318y.$$

The PSF coeffs are the parameters A_5, A_6, A_7 described in section 3.2.1 in the main part of this thesis. The numbers Sig2x, Sig2y are the initial guesses for A_5 and A_7 . Also listed is the initial guess for the sky and the size of the box used for profile fitting and aperture photometry.

After the file header there follows a table looking like this:

N	T	Sky	X	Y	Int	Sig2x	Sigxy	Sig2y
1	1	2219.91	385.13	179.26	11854.16	3.7005	-0.0009	3.9811

2	4	2270.87	338.10	215.47	7363.22	3.7005	-0.0009	3.9811
3	4	137.37	9.32	218.81	14224.61	4.6521	0.0022	4.2549
4	3	2322.54	338.96	231.48	10935.57	3.7005	-0.0009	3.9811
5	8	2896.87	357.00	229.00	13720.00	0.9000	-1.0000	0.9000

StrAp	SkyAp	Fit	Apr	G
4.1154E+05	2098.29	0.003	0.002	0
3.9261E+05	2114.33	0.021	0.002	0
0.0000E+00	333.47	0.056	0.000	0
4.5185E+05	2189.32	0.021	0.002	0
0.0000E+00	2193.25	0.000	0.000	0

The first column, N, is the star number. The type of object, T, in the second column is of particular importance, and has the meanings explained in Table 3.1 of the main part of this thesis. Next is the fitted sky value and the coordinates of the star. The intensity as given by A_4 parameter is in the column labeled Int. The columns Sig2x, Sigxy, Sig2y are the values of the shape parameters A_5, A_6, A_7 respectively. The total sky-subtracted electron count in the aperture is given in column StrAp while SkyAp is the number of sky electrons per pixel. Fit and Apr are approximate errors in the PSF fitting and aperture magnitudes. The last column, G, is not used.

A.4.3 The contents of “ccdpick.out” file

The output after running ccdpick is stored in file called ccdpick.out, the contents of this file might look like this:

```
1
2455506.40498    1   -5.4541   -8.6044    1.4862
2455506.41374    2   -5.5128   -8.6121    1.4549
...
2455508.38105   17   -5.4708   -7.2747    0.7985
2455508.41044   18   -5.7309   -7.6317    0.8418
    -1.00000     1    0.0000    0.0000    0.0000
2
2455506.40498    1   -4.6518   -7.8252    1.4862
2455506.41374    2   -4.7110   -7.8269    1.4549
...
2455508.41044   18   -4.8877   -6.8285    0.8418
2455508.44372   19   -6.0445   99.0000    0.7891
    -1.00000     1    0.0000    0.0000    0.0000
```

The first column is the HJD, this is followed by the running number, the profile fitted magnitude, the aperture photometry magnitude and the measure of the size of the stellar image. The measurements of each star are terminated by a line containing -1 for the HJD.

A.5 The 1.0-m telescope

The 1.0-m SAAO telescope (or Elizabeth telescope) is an optical telescope which is mounted equatorially, equatorial mount is the one where the telescope can move in North-South direction on declination (δ) and East-West direction on right ascension (α). This mounting is such that the axis around which the telescope is rotating/moving in α is parallel to Earth's rotation axis, meaning that this telescope axis points towards North and South celestial poles. 1.0-m telescope has the f -ratio given by $f/16$ which was set when the telescope was moved to Sutherland site from Cape Town. This telescope was built in 1964 by Grubb Parsons with the f -ratio of $f/20$. Mostly, the telescope is used for CCD photometry, the instruments used for science are SAAO CCD Camera (or STE4) and UCT Photometer (or STE3). The details about STE4 are given in the main part of this thesis and STE3 properties are given below. Fig. A.12 shows the external appearance of the telescope structure (Dome) and the optical diagram of the telescope.

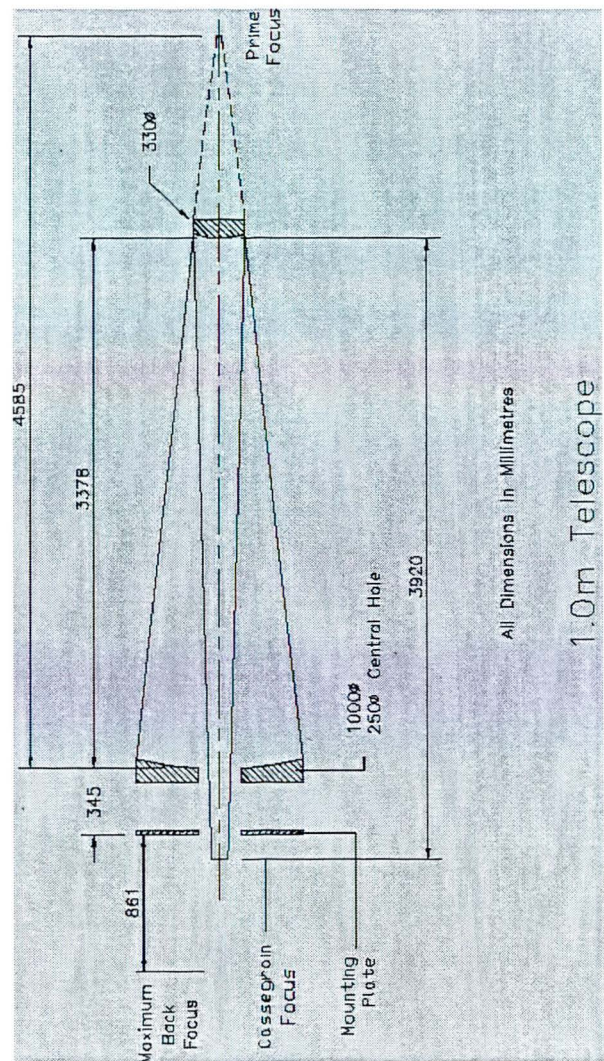
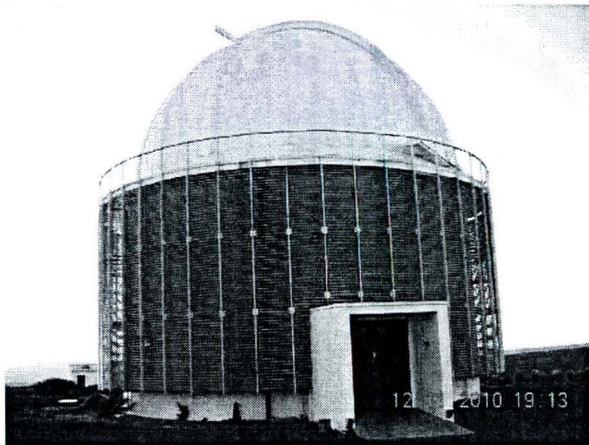


Figure A.12: A figure showing the external structure (left) of the 1.0-m telescope and its optical diagram (right) showing how the light beam is being reflected until it reaches the detector.

The 1.0-m telescope was designed such that its building puts the limits on the view of the sky, meaning the telescope can access observed target when is at the certain altitude, otherwise the building structures obscure the view of the sky. The declination limits of the telescope at the certain hour angles (HA) on the East and West are given on Table A.1.

DEC	HA East	HA West
-80	2:55	5:00
-70	3:50	5:00
-60	4:10	5:00
-50	4:15	4:45
-40	4:05	4:30
-30	3:55	4:15
-20	3:40	3:55
-10	3:25	3:35
0	3:00	3:05
+10	2:30	2:35
+20	1:40	1:50
+30	0:20	0:00

Table A.1: A table showing the limits, in declination, of the 1.0-m telescope ([http://www. saao.ac.za/facilities/telescope/1.0-m/](http://www.sao.ac.za/facilities/telescope/1.0-m/)).

As it was already mentioned, 1.0-m telescope has the camera for guiding it to follow the field it is pointed to, this is Acquisition & Autoguiding Camera (A/G Camera). The internal mechanism of the box within which the beam is directed to A/G camera is shown on Fig. A.13.

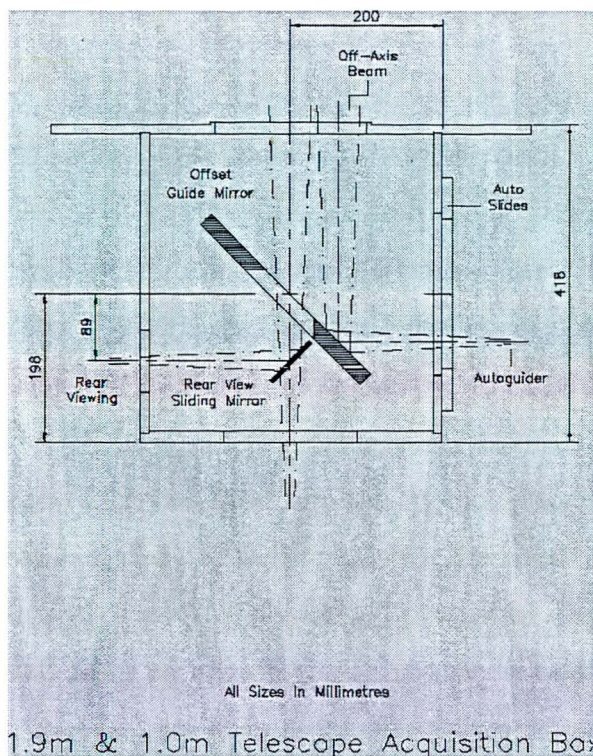


Figure A.13: A figure showing the Acquisition Box of the 1.0-m telescope within which the light can either be allowed to pass to the science detector or be reflected to the A/G camera. (<http://www.sao.ac.za/facilities/telescope/1.0-m/>).

The internal structure of the telescope is orientated as follows: the building has two floors, the ground floor houses the kitchen, toilet and the maintenance rooms and the first floor has two small rooms. One room is for storing computers which are running other systems of the telescope, and the other room is the observer's room (or the warm room as it is normally called). The warm room contains three computers, two of them are responsible for telescope status and controlling the science camera, respectively. The computer monitoring the status of the telescope works with the A/G camera for guiding the telescope to the desired field during observations. The third computer is mainly used for saving the observed data from the computer controlling science camera, and this third computer can also be used for internet purposes and for doing the data reductions while observing. Fig. A.14 shows the picture of the interior North-side of the warm room. The following paragraph gives the properties of STE3.

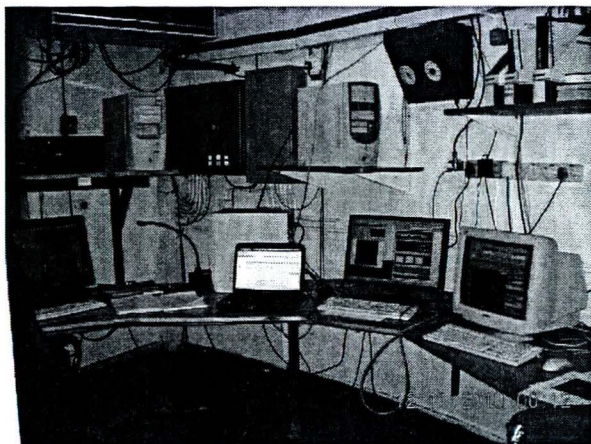


Figure A.14: An image showing the interior appearance of the warm room in the 1.0-m telescope.

The UCT CCD Photometer (or STE3) is a “Wright Instruments Peltier-cooled camera”. It is used in 1.0-m, 1.9-m and also in 0.75-m telescopes in frame-transfer mode for doing high-speed photometry of short period variables. The Table A.2 below shows the properties of STE3 of which some are specific to 1.0-m telescope.

No. of pixels	576×420
Pixel size	22 μm
Dark current	0.05 $\text{e}^-/\text{pix/s}$
Readout noise	10 e^-/pix
Gain	1 = 10 e^-/ADU
ADC max	32768 ADU (16 bit)
CCD saturates	250 000 e^-
Plate Scale	77 $\mu\text{m}/\text{arcsec}$ (1.0-m)
22 μm pixel size	0.28 arcsec (1.0-m)
1 arcsec seeing prebinning	(2×2) (1.0-m)
Number of pixels	190×130 (1.0-m)
Readout time	2 sec (1.0-m)
Field of view	109×74 arcsec (1.0-m)

Table A.2: A table showing the properties of the UCT CCD Photometer, some of the properties are specific when the Photometer is mounted to the 1.0-m telescope and they are shown with “1.0-m” in brackets (<http://www.sao.ac.za/facilities/instruments/uct-ccd/>).

A.5.1 Setting the zero points

When using the 1.0-m telescope, it is important to set the zero points for the field within which your target is located. This enhances the pointing of the telescope and makes it easy to acquire the correct field when pointing the telescope to your target. The stepwise procedure for setting the zero points is given below, and it must be noted that all highlighted words are buttons to be clicked on the interface of the Telescope Control System (TCS), this interface is shown on Fig. A.15 below.

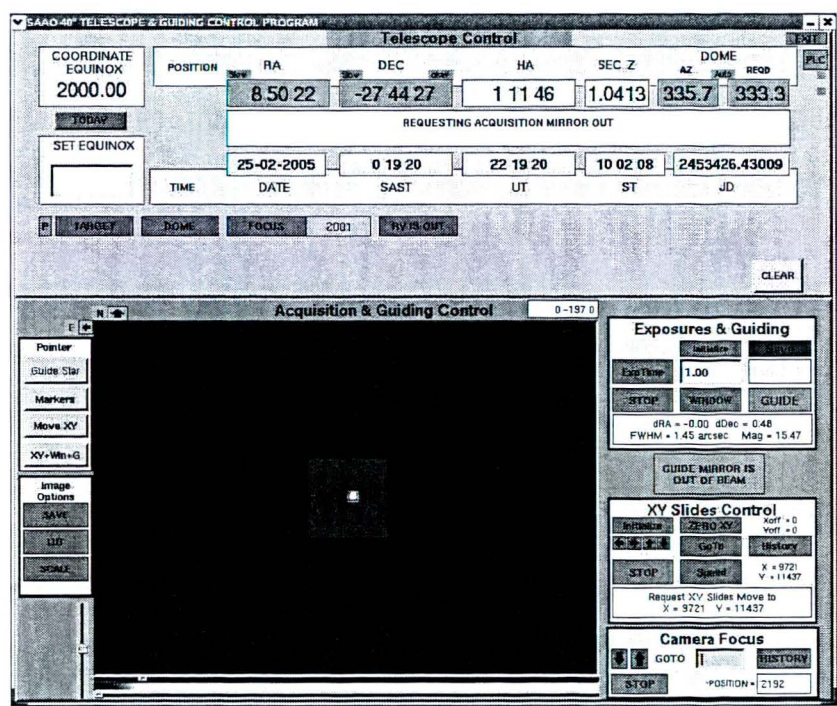


Figure A.15: A figure showing the interface of the TCS for controlling and monitoring the status of the telescope during the observations (Credit: Potter S. & Worters H.; TCS user’s guide; version 3; 2010).

Looking at the TCS interface on Fig. A.15, the zero points setting procedure is as follows:

- Click the **Initialize** button in “XY Sides control”.
- By using Sutherland Almanac, which is available on the books shelf inside the warm room, select a bright star near your science target. This can be selected

by looking on the bright stars with RA and DEC values closer to your target's coordinates and for the brightness of the star, one should

select the star which is brighter than $V = 5$ -mag to make it easy to see it through the finderscope of the telescope.

- Type in the "COORDINATE EQUINOX" of the selected bright star coordinates and press ENTER on the keyboard. The coordinates equinox can be read from the top of the page containing the bright star.
- With **GUIDE MIRROR** in beam, slew to and acquire the bright star. To put the Guide Mirror in beam, click on **GUIDE MIRROR IS OUT OF BEAM** button, which will be green and it must turn red when the Guide Mirror is in beam, with the message: **GUIDE MIRROR IS IN BEAM**.
- Set **WINDOW** to "Full Frame" by clicking **WINDOW** button in "Exposures & Guiding" and select "Full Frame", with **Exp Time** ≈ 0.05 sec also in "Exposures & Guiding" place the bright star at the center of the image display area. This is done by looking through the finderscope attached to the telescope and use the hand-set, which is also attached to the telescope, to place bright star at the center of the red-cross shown when looking through the eyepiece of the finderscope. If the red-cross is not shown, then switch it on with flip-switch located on the left-hand side, just above the eyepiece of the finderscope.
- Click **P** then click and hold **ZERO POINTS** and in "Enter RA & DEC of Star" type the coordinates of the bright star which is centered and click **ENTER**.
- The zero points for RA & DEC will be set and displayed in the white message box. One must note these zero point values down for future use for that part of the sky, because different parts of the sky require different zero points. These noted zero points will need to be re-entered in case TCS needs to be re-started.
- To use previously determined zero points, click **P** then click and hold **ZERO POINTS** and select "Enter zero points directly", this will provide space for

typing previously determined zero points. After entering them click **SET** and **CLEAR**

After setting the zero points, the most important thing to do is to apply them to the coordinates of your target. Before doing so, ensure that the COORDINATE EQUINOX is set back to that of your target coordinates. The steps on how to get the corrected coordinates for your target by applying zero points are explained below:

- Click **O** which was shown during zero points setting when **P** was clicked, this should show **P** again, then click **TARGET** and type in the RA & DEC of your target, then click **SET** and **CLEAR**.
- Confirm that “COORDINATE EQUINOX” is set to that of your target coordinates, if that is correct then slew the telescope to target coordinates which are displayed on the monitor, in orange, above the coordinates written in black on the monitor on observing floor, near the main telescope control hand-set.
- To check whether the field pointed with the telescope is correct, move **GUIDE MIRROR OUT OF BEAM** to allow the beam to the science detector, this should turn green when the Guide Mirror is out of beam.
- Take a 10s **Snapshot** on science detector computer and compare the readout image with your finding chart. In order to make printed finding chart similar to the readout image, look at the finding chart from the back, you will need the light in front of the chart to see the image. Otherwise the finding chart and the image on the science CCD computer screen looks like on Fig. A.16 compared to the finding chart.

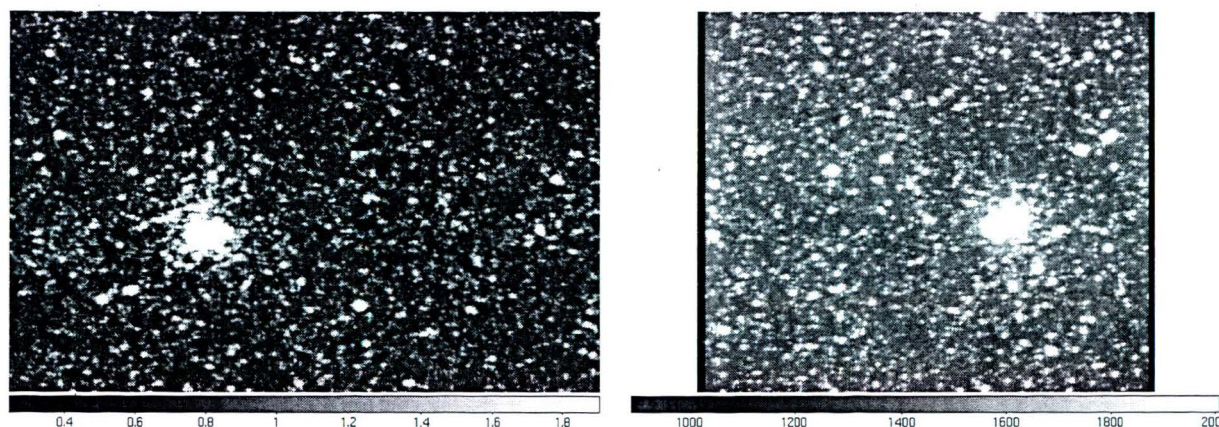


Figure A.16: A figure showing how the finding chart (left) of the observed target compares to the image readout (right) from the science CCD camera. It can be seen from the images that left-right of one image is right-left of the other and top-bottom of one image is bottom-top of the other image.

The above steps for setting zero points and applying them to your target coordinates should make it easy to acquire the correct field for observed target. As mentioned above, different parts of the sky require different zero points.

(Credit: Potter S. & Worters H.; 1.0-m Telescope Control Software (A User's Guide); Version 3; 2010.)