Spectral Results for the Blue Plume Stars in Canis Major Overdensity

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SPECTRAL RESULTS FOR THE BLUE PLUME STARS IN CANIS MAJOR OVERDENSITY

THESIS

A thesis submitted in partial fulfillment of the requirements for the degree of Master of Science in Physics in the College of Arts and Sciences at the University of Kentucky

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ABSTRACT OF THESIS

SPECTRAL RESULTS FOR THE BLUE PLUME STARS IN CANIS MAJOR OVERDENSITY

We present distances and kinematics and look at the possible populations for the blue plume (BP) stars in the Canis Major Overdensity (CMa). We conducted a medium resolution spectral survey on the BP stars (N=303) in CMa (centered at \( l = 238^\circ ; b = -8^\circ \)) using the data from AAOmega Spectrograph. We used a modified version of the Statistics-sensitive Non-linear Iterative Peak-clipping (SNIP) algorithm to normalize our fluxed absorption spectra. After determining the radial velocities from measurements of strong absorption features for the stars we use a Bayesian analysis of spectral feature strengths and photometric colors to determine Teff, Logg and [Fe/H]. Our procedure makes use of grid for model synthetic spectra computed using SPECTRUM with Atlas9 model atmospheres and Kurucz model colors. We determine the absolute magnitude using the stellar parameters and BaSTI isochrones and compute distances and ages for the BP stars.

Our analysis of the BP stars indicates Teff ranging from 6500K to 8000K, metallicity ranging from 0.0 to -1.0 with an average of -0.5. We found for this temperature range that the surface gravity of the stars could not be well constrained. From the spatial and kinematics results we found that most of the stars are thick disk stars with a small mixture of thin disk stars. The stars are most likely a mixture of thick disk blue stragglers and normal A-type stars preferentially seen to greater depths due to the low dust extinction in this location of the Galaxy.

KEYWORDS: Blue Plume (BP) stars, Canis Major Overdensity (CMa), Bayesian analysis.

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Chapter 1 - Introduction

1.1 Motivation:
The spatial, kinematic and metallicity characteristics of the outer disk of the Milky Way galaxy has been a mystery due to a complex dynamic history that includes warps, flaring and potential perturbations from merging dwarf galaxies. An interesting location in the disk is the low dust extinction area centered at Galactic coordinates $l = 238^\circ$ and $b = -8^\circ$ which contains a large number of spectral-type A/B stars, popularized as the Blue Plume (BP) stars. These stars are thought be hot, young, main-sequence stars that inhabit the blue part of the color magnitude diagram.

Canis Major Overdensity (CMa) was discovered by Martin et al. (2004) using data from the 2MASS (the two micron all sky survey) infrared survey. The authors found a significantly high density of M-giant stars (Martin et al. 2004) at this location, and further spectroscopic follow-up (Martin et al. 2005) of the M-giants and red clump stars in the region, suggested that CMa may be a spatially and kinematic distinct population from the Galactic disk. With a size of $\sim 2$ kpcs the CMa was construed as a merging dwarf galaxy remnant. The aforementioned BP stars were initially thought to be a part of this dwarf galaxy as argued by Bellazzini et al. (2004); Dinescu et al. (2005); Martínez-Delgado et al. (2005).

Later work by Mateu et al. (2009) showed no evidence of an ancient population in the CMa bringing the entire notion of a galaxy at this location into serious doubt. Carraro et al. (2005), Moitinho (2006) and Vázquez et al. (2008) showed the location of the CMa stars are consistent with the location of the Orion Spur and the outer Norma-Cygnus spiral arm. Although it is currently not well accepted that the Canis Dwarf
galaxy exists, the BP stars can still be very useful in helping to unveiling the true nature of the CMa and the characteristics of the outer disk of the Galaxy.

We explored these BP stars with spectroscopy to better constrain the spatial, kinematic, and metallicity distribution in the outer part of the disk of Milky Way and to gain insight into the Galactic disk evolution.

In the next few sections I will explain some important astronomical concepts that we use to determine stellar parameters for the BP sample.

1.2 Photometry:

The science of measuring and calibrating the apparent brightness of a star is called photometry. Astronomers measure the apparent brightness of a star using a telescope with an attached light-sensitive instrument. Although it sounds like a simple task: counting the number of photons arriving from the star and we are done, but it is one of the most difficult techniques in astronomy.

The apparent brightness of a star’s light or the energy flux reaching us from a star can be stated as:

\[ F = \frac{L}{4\pi d^2} \]  

(1.2.1)

Where \( L \) is the luminosity of the star, in W, or the total energy output of the star per second, and \( d \) is the distance to the star from us, in meters. Apparent brightness is measured in watts per square meter (W/m\(^2\)).

Astronomers commonly express a star’s apparent brightness in terms of apparent magnitude (\( m \)), and the star’s luminosity in terms of absolute magnitude (\( M \)). The
absolute magnitude of a star is defined as the magnitude the star would have if we placed at a distance of 10 parsecs (pc) or 32.6 light years from the observer, assuming no astronomical extinction of starlight.

If we want to compare two stars that have different know fluxes, \( F_1 \) and \( F_2 \), we can use the following equation:

\[
m_2 - m_1 = 2.5 \log_{10} \left( \frac{F_1}{F_2} \right)
\]  

(1.2.2)

This equation is also useful for relating two different magnitude stars in terms of their relative fluxes.

Substituting equation (1.2.1) into equation (1.2.2) a relation between a star’s apparent magnitude and absolute magnitude can be obtained for a given distance \( d \) of the star from the Earth:

\[
m - M = 5 \log_{10} d - 5
\]  

(1.2.3)

In this expression \( m - M \) is called as the distance modulus of the star.

### 1.3 Filters:

In the equation (1.2.2) for magnitude, one of the fluxes is the standardizing flux and another one might be the total (all wavelengths or ‘bolometric’) light flux, or a monochromatic flux density \( F_\lambda \), or the flux in a particular range of wavelengths. Much of astronomy is concerned with the flux through the five standard U, B, V, R, and I filters. The characteristics of these filters are detailed in Figure 1.1, but essentially each filter allows transmission of a range of wavelengths \( \Delta \lambda \), centered on a particular
Figure 1.1: The characteristics of the standard filters: U (ultraviolet), B (blue), V (visible), R (red), I (infrared) from Allen’s Astrophysical Quantities p. 387

The name of the filter reports the type of light allowed to pass: U (ultraviolet), B (blue), V (‘visible’, actually green), R (red), and I (infrared). The filters are broad enough that the resulting bin really doesn’t well represent a value for $F_{\lambda}$. The magnitude of a star as measured through a B filter is called the B magnitude and is simply denoted: B. We can now form color indices just by subtracting two magnitudes. For example, the most common color index is $B - V$:

$$B - V = 2.5(\log_{10}\left(\frac{F_V}{F_{V0}}\right) - \log_{10}\left(\frac{F_B}{F_{B0}}\right))$$

$$= 2.5(\log_{10}\left(\frac{F_V}{F_B}\right) + \log_{10}\left(\frac{F_{B0}}{F_{V0}}\right))$$

$$= 2.5 \log_{10}\left(\frac{F_V}{F_B}\right) + \text{Constant} \quad (1.2.4)$$
Where $F_{B0}$ and $F_{V0}$ are the standardizing flux through the B and V filters respectively. From the above equation we can see that $B - V$ is related to the flux ratio $F_V/F_B$ and so is an intrinsic property of the star related to temperature.

1.4 Interstellar Reddening:

The absorption and scattering (by interstellar gas and dust) of electromagnetic radiation emitted from an astronomical object is termed as interstellar extinction. Interstellar reddening is a phenomenon associated with interstellar extinction where the spectrum of electromagnetic radiation from an astronomical object changes characteristics from the originally emitted spectrum. The extinction due to interstellar dust and gas is not equally effective at all wavelengths. The shorter the wavelength, the higher the extinction, which means the blue light is scattered and absorbed more strongly since the dust grains in the interstellar medium have a typical size that is comparable to the wavelength of blue light. Therefore, objects behind a large column of dust appear redder than they really are. This effect is known as interstellar reddening and must be taken into account when determining temperatures and distances for stars.

We can determine the degree of reddening by measuring the $B - V$ color index of the object and compare that to its true or intrinsic color index $(B - V)_0$ using the following equation:

$$E(B - V) = (B - V) - (B - V)_0 = A_B - A_V$$  \hspace{1cm} (1.3.1)

Where $A_B$ and $A_V$ are the extinctions in the blue and visual bands respectively.
Since both interstellar reddening and extinction are result of the interaction between starlight and interstellar dust and gas, it is reasonable to expect that the more reddening there is, the more extinction there will be. In fact, studies have shown the extinction and reddening are related by:

\[ A_V = R_V E(B - V) \]  

(1.3.2)

Where, \( R_V \approx 3.1 \). Often \( A_V \) is called as total extinction and \( E(B - V) \) is called as selective extinction.

In astronomy, interstellar reddening plays an important role in distance measurement. From equation (1.2.3), we get the equation for measuring the distance as:

\[ d = 10^{0.2(m-M+5)} \]

(1.3.3)

This equation represents the distance without the extinction. We need to brighten the apparent magnitude to account for the loss of light to get the correct distance. Since lower magnitudes are brighter, \( A_V \) needs to be subtracted from the apparent magnitude. The revised equation for distance (in parsecs) is thus:

\[ d = 10^{0.2(m-M-A_V+5)} \]

(1.3.4)

1.5 Spectroscopy:

Photometry gives us the apparent brightness of the stars we study and it also guides us to determine the surface temperatures and distances of the stars. To determine the other properties of a star, we have to take full advantage of the information encoded in a star’s light. Since the chemical composition of a star is more directly imprinted in
the spectra as well as the radial velocity through Doppler shifts in the spectral lines, analyzing star’s spectrum allows greater details to be discovered.

In our proposed method, we used spectra of the stars to estimate the metal abundances, [Fe/H] as well as the surface temperatures since H Balmer lines give us much greater accuracy due to the fact that they are not significantly affected by dust. It also has the potential to constrain surface gravity due Stark pressure broadening of the Balmer line wings.
2.1. Target Selection:

We have obtained medium resolution spectra (R ~ 2000) of the Blue Plume (BP) stars (N=303) at the center of CMa using the data from AAOmega Spectrograph supplied by our collaborator Kenneth Carrell. We used a window (3900 - 5250Å) of the visible light wavelengths to investigate the spectra of the stars since this region contains the Hδ, Hγ and Hβ lines as well as many useful metal-line. The spectroscopic targets were chosen from UBVRI photometry supplied by Giovanni Carraro from which we selected both BP and Red Clump stars to be observed (Fig 2.1). The data were reduced using standard techniques in IRAF by Kenneth Carrell.

![Color-magnitude diagram of the BP and Red Clump stars.](image)

Fig 2.1: Color-magnitude diagram of the BP and Red Clump stars.
2.2 Normalization:

We used a modified version of the Statistics-sensitive Non-linear Iterative Peak-clipping (SNIP) algorithm supplied by Kenneth Carrell to normalize the fluxed absorption spectra. Algorithm D from Morháč, M. (2009) is used and modified to be applicable to absorption spectra since they used the algorithm for emission spectra. This algorithm uses two parameters ‘smoothing’ and ‘simultaneous smoothing’ to smooth out noise from the spectra. Where ‘smoothing’ picks a single width from the entire continuum of the spectrum, smooth out that width and goes to the next width to perform the similar action. For example, if we choose a ‘smoothing’ value of 50, it means that anything less than ~100 pixels will be smoothed out in the background estimation. For both synthetic and observed data, a similar ‘smooth’ value is required to clip the noise from the spectra.

‘Simultaneous smoothing’ is a boxcar-like smoothing which smooth out the neighboring data points. Since the synthetic data do not have noise while the observed data do, therefore, ‘simultaneous smoothing’ is not required for the synthetic data.

We tested out the program for both hot and cool stars and for both synthetic and observed spectra. An example for finding the optimized ‘smoothing’ value (m) is shown in fig 2.2 where we examined two synthetic spectra with parameters Teff = 13000K, logg = 4.5, [Fe/H] = 0.0 and Teff = 8000K, logg = 4.5, [Fe/H] = 0.0 respectively. Although from the flux vs wavelength graph it seems that m = 42 is better for 8000K and m = 25 is better for 13000K because of the slight dips in the background in the Balmer lines for m = 25 in the 8000K spectrum but from the
normalized flux vs wavelength graph we see that $m = 25$ gives us a better normalization for both of the spectra.

Fig 2.2: Plots of synthetic spectra for both hot and cool temperature.
We used a ‘smoothing’ value of 25 for both synthetic and observed spectra and a ‘simultaneous smoothing’ value of 1 for observed spectra in order to have the fit reach to the continuum of the synthetic spectra. A few examples of how both of these parameters work for observed spectra are shown in the following figures:

Fig 2.3: (a) Normalizing an observed spectrum using smoothing value = 10 and simultaneous smoothing value = 1, (b) smoothing value = 15 and simultaneous smoothing value = 1.
Fig 2.3: (c) Normalizing an observed spectrum using smoothing value = 25 and simultaneous smoothing value = 1, (d) smoothing value = 35 and simultaneous smoothing value = 1.
Fig 2.3: (e) Normalizing an observed spectrum using smoothing value = 50 and simultaneous smoothing value = 1, (f) smoothing value = 75 and simultaneous smoothing value = 1.
2.3. Velocity Correction:

We determined the line-of-sight radial velocities using the weighted mean of line centers from strong absorption features and use this velocity to shift our spectra to rest frame velocities. I wrote a python code that calculates rest frame wavelength ($\lambda_0$) by using Doppler shift formula (equation 2.3.1) and shift our spectra to rest frame velocities:

$$\lambda_0 = \frac{\lambda}{1 + \frac{v}{c}}$$  \hspace{1cm} (2.3.1)

Where $\lambda$ is the shifted wavelength, $v$ is the velocity of the source measured along the line of sight, and $c$ is the speed of light ($3.0 \times 10^5$ km/s).

After shifting the observed spectra to rest frame velocities, we also interpolated both observed and model spectra so that we can subtract the flux values exactly at the same wavelength pixel by pixel while calculating the $\chi^2$-values in the later part of the analysis.

2.4 Statistical Analysis:

We compare the observed spectra with the model spectra in order to find the best-fit model spectrum to each observed spectrum. I wrote a python script that calculates the $\chi^2$- values for the H Blamer lines (H-δ, H-γ, and H-β) and the metal regions for each of the stars from a comparison with a grid of model synthetic spectra (computed using SPECTRUM (Richard Gray) with Atlas9 Kurucz ODF model atmospheres). The wavelength ranges I used for different regions of the spectra are: 4060Å - 4140Å for H-δ, 4300Å - 4380Å for H-γ, 4820Å - 4900Å for H-β, 4140Å - 4300Å for metal...
region one, 4380Å - 4820Å for metal region two, 4900Å - 5216Å for metal region three. The chi square ($\chi^2$) analysis is done through a pixel-by-pixel comparison of the spectrum of a given star against 2128 model spectra. The model spectra cover a temperature range of 6000K - 19000K with a separation of 250K from 6000K to 13000K and a separation of 1000K from 13000K to 19000K. For each of the temperatures, the range of logg values is from 1.0 to 5.0 with a separation of 0.5 and the values for metal abundances are 0.0, 0.2, 0.5, -0.5, -1.0, -1.5, -2.5 and -2.5.

The formula we used to calculate the $\chi^2$-values is:

$$\chi^2 = \sum_{n=1}^{N} \left( \frac{f_m - f_o}{\sigma_n} \right)^2$$  \hspace{1cm} (2.4.1)

Where $f_m$ and $f_o$ are the normalized fluxes of model star and observed star in a corresponding pixel respectively. The value of $\sigma$ is calculated by using the formula:

$$\sigma = \sqrt{\frac{2}{|\text{flux}| \times \text{gain}}}$$ \hspace{1cm} (2.4.2)

Where the value of $\text{flux}$ is the original flux value before normalization for every pixel and we used a value of 3.36 e'/ADU for $\text{gain}$. 
The following figures show a comparison of an observed spectrum with a model spectrum for both low and high $\chi^2$-values:

![Graph 1](image1.png)

![Graph 2](image2.png)

Fig 2.4: An observed spectrum with a model spectrum for low (a) and high (b) $\chi^2$-values.
If we take a closer look at the H-δ absorption line of the above figures, we see the following:

Fig 2.5: H-δ absorption line for an observed spectrum with a model spectrum for low (a) and high (b) $\chi^2$-values.
Figure 2.6 shows a zoomed in view of the variation in metal abundance for different metallicities at a given temperature and surface gravity. The best-fit line corresponds to the lowest $\chi^2$-value for metal region two (4380Å - 4820Å) that represents the average variation of the metal abundance from the observed spectrum. We observe a few spikes in the image and those come from the normalization fit of the observed spectrum that goes up a little higher.

Since we calculated the $\chi^2$-values separately for different regions of a spectrum, not simultaneously for the whole spectrum, we used a reduced $\chi^2$ method to calculate the likelihoods for each portion of the spectrum. It also helps us to evaluate the $\chi^2$-values for the H – Balmer lines and the metal regions on the same scale since H – Balmer
lines are in a very small wavelength range compared to the metal regions. The reduced $\chi^2$ is defined by:

$$\chi^2_{\text{red}} = \frac{\chi^2}{K} \quad (2.4.3)$$

Where $K$ is the number of degrees of freedom and the degrees of freedom is defined as $(N - P)$, $N$ is the number of data points (or the number of pixels in our case) and $P$ is the number of parameters in the fitting function that is being used. In our case, we have three parameters – effective temperature, surface gravity and metal abundances.

I wrote another code in python to calculate the likelihoods for each $\chi^2_{\text{red}}$ value where the likelihood function is defined as:

$$L = \frac{1}{\sqrt{(2\pi)^K|\Sigma|}} e^{-\frac{\chi^2_{\text{red}}}{2}} \quad (2.4.4)$$

Where $K$ is the number of degrees of freedom and $\Sigma$ is the average variance of a particular region of a spectrum.

The python script I wrote for calculating the $\chi^2$-values for the different regions of a spectrum also calculates the pixel-by-pixel variances that I used to calculate the average variance for a particular region of spectrum so that I can use it to calculate the likelihood of that region.

We found the probabilities for the stellar parameters by multiplying the individual likelihood for different regions and used it to calculate the expectation values and the standard deviations for different parameters of the stars. The formulas we used to calculate the expectation values and the standard deviations are:
We found degeneracies in the H – Balmer line temperature while we were examining the stellar parameters with their probabilities. This degeneracy is the result of the Balmer line strength reaching a maximum value near surface temperatures of $T_{\text{eff}} \approx 8750$ K. The diminishing strength of the Balmer lines beyond this temperature results in two possible temperature values for a given Balmer line strength. To avoid this degeneracy, which can occur in our sample spectral range, we reddening corrected our $B - V$ colors using Schlafly & Finkbeiner reddening 2011 (ApJ 737, 103) and constrained the temperatures for our stars using a Kurucz model color calibration. For the reddening correction of $B - V$ we rewrite equation (1.3.1) as:

$$ (B - V)_0 = (B - V) - E(B - V) $$

If we look at the center of our field (Galactic coordinates $l = 238^\circ$ and $b = -8^\circ$) on the Schlegel dust maps (Fig. 2.7 (a), image size 10.0 degrees) we see the dust emission is quite low compared to the other regions at this location.

The dust emission looks higher if we take a closer look (Fig. 2.7 (b), image size 2.0 degrees) in the same image, but this is basically due to the changes in the contrast in the image. Since we do not have the higher dust emission region in the second image, which we had in the first image, it adjusts the contrast in the second image.
We looked up the reddening distribution in this field and according to Schlafly & Finkbeiner reddening, 2011, the maximum reddening is 0.2855 and the minimum reddening is 0.2351 with a mean value of 0.2574 ± 0.0142. That means the difference across the field is ~0.05 and which implies that the reddening in our field does not change drastically.

![Schlegel dust maps](image)

Fig 2.7: Schlegel dust maps for our field (a) image size = 10.0 degrees, (b) image size = 2.0 degrees.

The reddening values we used for our stars are average interpolated values over a pixel. It is because the resolution the Schlegel dust maps is low and that makes the size of a pixel much larger than the size of a star. If the reddening is differentially changing on scales smaller than the resolution element then the interpolation is uncertain. But if the amount of change is small across the field then the error in the interpolated value will also be small.

There were some RR Lyrae stars in our observed sample and we found them while we were doing the reddening correction. Since these stars are pulsating stars therefore we do not have the correct radial velocities and we neglected these stars. We also got rid of the logg values of 0.0 and 0.5 from our grid.
Since our goal is to set a prior to the temperature range for each of the stars so that we can remove the degeneracy, we need to know the maximum and minimum values of $B - V$ colors from the model grid. If we know maximum and minimum values of $B - V$ colors we can find the upper and lower cut off values for temperature of a given star using the reddening corrected $B - V$ values via interpolation.

The following graph shows maximum and minimum values of $(B - V)_{\text{model}}$ with their corresponding temperatures:

![Fig 2.8: Temperature vs $(B - V)_{\text{model}}$ graph](image)

Once we found the priors for the temperatures for each of the stars, we calculated the $\chi^2$-values only for the constrained temperatures and following the similar steps as described previously we calculated the expectation values and the standard deviations using equations (2.4.5) and (2.4.6) for the stellar parameters.
Chapter 3: Results

Our project was to develop a new method for doing spectral analysis on the Canis Major Over-density’s (CMa) Blue Plume (BP). We used $\chi^2$ analysis to find the likelihoods for the stellar parameters and used Bayesian analysis based on a posterior distribution of the stellar parameters by constraining the temperature ranges for each of the observed stars using Schlafly & Finkbeiner reddening 2011 (ApJ 737, 103). The prior distribution of the temperatures helps us to avoid degeneracies in the temperatures and provides us an acceptable range of temperatures to calculate the expectations values of the stellar parameters. The following figure shows a frequency histogram of the number of the stars versus the effective temperature:

![Fig 3.1: A frequency histogram of the number of the stars versus the effective temperature](image)

Fig 3.1: A frequency histogram of the number of the stars versus the effective temperature
We looked at the cooler ($\leq 8500$ K) stars of the sample for which we can currently get more reliable metal abundances i.e. we observed smaller error ($<0.6$) in the metal abundance for cooler stars. As the temperature gets higher for a star the metal lines become too weak to constrain them very well, this causes a higher error in the metal abundance of the star. The following figure shows the frequency histogram of the number of stars versus the metal abundance for the cooler stars:

![Frequency Histogram](image_url)

*Fig 3.2: A frequency histogram of the number of stars versus the metal abundance for the cooler stars ($\leq 8500$K)*
We determine the absolute magnitude using the stellar parameters and BaSTI isochrones and a frequency histogram of the number of the stars versus the absolute magnitude is shown in figure 3.3:

![Figure 3.3](image_url)

**Fig 3.3**: A frequency histogram of the number of the stars versus the absolute magnitude

We get the V color magnitude from private communication with Giovanni Carraro to find out the distance of the stars. We corrected V values using the formula:

\[
V_0 = V - 3.1 \, E(B - V)
\]

(3.1.1)

Where we again used the reddening values for the stars from Schlafly & Finkbeiner reddening 2011 (ApJ 737, 103).
We calculated the distances using the absolute magnitudes ($m_v$) and the corrected $V$ values by using the following equation:

$$d = 10^{\frac{V_0 - m_v + 5}{5}} \quad (3.1.2)$$

The following figure shows the frequency histogram of the number of Stars versus their distances:

![Frequency histogram figure](image)

**Fig 3.4**: A frequency histogram of the number of Stars versus distance (pcs)

We calculated galactic standard of rest line-of-sight velocity ($v_{GSR}$) to find the location of the stars in the galaxy by using equation (3.1.2):

$$v_{GSR} = v_{hetio} + 9 \cos(l) \cos(b) + 232 \sin(l) \cos(b) + 7 \sin(b) \quad (3.1.2)$$
Where $v_{\text{helio}}$ is the heliocentric correction to the radial velocity ($v_{\text{rad}}$) and we determined this by adding 3.6 km/s to the radial velocities of the stars.

We plot distance versus galactic standard of rest line-of-sight velocity ($v_{\text{GSR}}$) graph with the model and the stars seem to be a mixture of thick and thin disk stars. We observed that the biggest clump is right at the thick disk that is an indication that the majority of the stars are thick disk stars. The plot is shown below:

![Figure 3.5: A plot of distance versus galactic standard of rest velocity ($v_{\text{GSR}}$) of the stars](image)

We plot logg versus Teff along with the model and we get a graph that is shown in figure (3.6). In this graph we observe some discrepancy with the model lines. This is may be because of the reddening values we used here which were expected to be higher than the used values.
Fig 3.6: A plot of logg versus effective temperatures of the stars

In our current analysis we find the majority of our stars have thick disk-like kinematics and metal abundances. The stars are well distributed along the line-of-sight with distances that consistent with the local Orion spur.
Chapter 4: Conclusions and Recommendations

Our analysis of combining spectroscopy with the reddening corrected photometry yields several parameters that help to explain the populations of stars in CMa that we observed. Implementing Bayesian analysis along with the $\chi^2$ analysis helps us to obtain an acceptable range of effective temperatures for the stars. It provides us a good way of examining the probability distribution for a given parameter so that we can constrain the parameter very well. It also opens up the opportunity to bring in additional observables (e.g. infrared photometry or ultraviolet photometry) to our analysis that can help us to improve our measurement of the stellar parameters.

The spectra combined with the reddening corrected photometry tell us about the distance and the kinematics of the stars. The stars seem to be a mixture of thick and thin disk stars and we find the majority of our stars have thick disk-like kinematics and metal abundances.

The metal abundance distribution ($[\text{Fe/H}] \approx -0.5$) is consistent with local (what we measure closer to the sun) thick disk metallicity distribution. In general, from the kinematics and metallicity distribution of the stars we can conclude that most of the stars are thick disk stars. The age of thick disk (~12Gyrs) and the type of the observed stars (Spectral type A stars) also tell us they are most likely to be blue stragglers (BSS).

The metallicity distribution function also reflects that there are lower abundances ($[\text{Fe/H}] \approx -0.5$) beyond the galactic radius of 10 kpcs i.e. the thin disk metallicity
becomes \([\text{Fe/H}] \approx -0.5\) at this distance from the galactic center. In this case, we could have some thin disk stars in our population.

The spectra can also tell about the surface gravity of the stars. Currently we have assumed that all the stars are main sequence stars (Appendix A) having a surface gravity of 4.0 but we would like to improve our surface gravity determination to avoid this assumption.

We observe that as the stars get hotter (\(\text{Teff} > 8500\text{K}\)) the error in metal abundance becomes higher. Therefore, we need to constrain the metal abundance at higher temperature in an effective way.

One way to extend metallicity measurements to higher effective temperature values is to make use of the CaIIK line, which is present in our spectra and remains measurable to \(\text{Teff} \sim 10000\text{K}\). We did not use this line in the current study because at low galactic latitudes there can be significant contamination from the interstellar Ca. This, however, can be modeled (Beers, 1990) and we intend to do this to be able to extend our metallicity coverage.

We also need to test the reddening values (Schlafly & Finkbeiner reddening 2011) that we used here is correct. Our data suggests that a somewhat higher reddening value for the majority of our stars is needed. (Appendix A)

It would also be interesting to look at the Red Clump (RC) stars (Figure 2.1). The RC stars are located at the same location with the BP stars. It would be interesting to do the same type of analysis on the RC stars and observe if they show the similar characteristics (distance, kinematics, metallicity) as the BP stars.
Appendix A

We have tested the reddening values for our stars using the $B - V$ (not reddening corrected) values and the equivalent width of the H – Balmer lines and we obtained the following graphs:

Fig A1: H – Balmer line Equivalent width versus B – V graph for different reddening values.

The red lines in the plots are for different surface gravity values ranging from 2.5 to 4.5 with a separation of 0.5 from the bottom line to the top line for a fixed temperature value of 8500K. We used SPECTRUM model generated from Kurucz model for the model equivalent width of the H – Balmer lines. Since the size of the H – Balmer lines does not depend on the reddening and reddening is only in the
horizontal direction we moved the model lines horizontally so that they fall on top of the equivalent width of the H – Balmer lines for the observed stars. The average amount we had to move is 0.325, which is a measure of the interstellar reddening for our stars.

Our measured reddening is significantly higher than the average reddening (~0.19) obtained from Schlafly & Finkbeiner reddening 2011 for our stars.

Figure A1 also gives us an idea about the surface gravity of the stars. If we look at the surface gravity values in figure A1, the best-fit model line to the observed stars corresponds to the surface gravity of 4.0, which is an indication that most of our stars are main sequence stars.

Another way of supporting our assumption that most of the observed stars are main sequence stars is to look at the following graph:

![Graph](image)

Fig A2: H – Balmer line Equivalent width versus B – V graph for a fixed reddening value with theoretical isochrones.
We used theoretical isochrones for a reddening value of 0.325 on this H – Balmer line Equivalent width versus B – V (not reddening corrected) graph and our stars fit well with the main sequence isochrones. The pink dashed line represents the horizontal branch line and it cuts across at the tail of the distribution and the stars under this line could be horizontal branch stars and they are very few in numbers.

We took the average reddening value (~0.19) for our stars from the Schlafly & Finkbeiner reddening 2011 and our estimated average reddening value (~0.325) to calculate the reddening corrected V magnitude so that we can compare the distances for our stars for this two reddening values. We found ~15% difference between the distances for these two reddening values and our estimated reddening value brings the stars closer than we calculated by ~500 pcs. This suggests that we can expect a systematic error of ~15% in our distances.
Appendix B

When we first began, we wanted to calibrate the reddening for our method. I went on an observing run at McDonald Observatory, Fort Davis, TX and used the 2.1m telescope to obtain low resolution spectroscopy of spectral type A/B stars in six nearby open clusters to calibrate the procedure of estimating reddening. We wanted to try to get the reddening for these stars because there is well determined reddening for them.

The following table shows the list of the clusters with the number of the observed stars from each cluster. The table also includes the reddening values for the clusters along with their distances and ages.

<table>
<thead>
<tr>
<th>Cluster Name</th>
<th>No. of Stars</th>
<th>E(B – V)</th>
<th>Distance (pcs)</th>
<th>Age (Myrs)</th>
<th>Other Name</th>
</tr>
</thead>
<tbody>
<tr>
<td>IC_4665</td>
<td>8</td>
<td>0.174</td>
<td>352</td>
<td>43</td>
<td>Collinder 349/ Mellote 179</td>
</tr>
<tr>
<td>NGC_225</td>
<td>5</td>
<td>0.274</td>
<td>657</td>
<td>130</td>
<td>-</td>
</tr>
<tr>
<td>NGC_6494</td>
<td>26</td>
<td>0.356</td>
<td>628</td>
<td>300</td>
<td>M23</td>
</tr>
<tr>
<td>NGC_6823</td>
<td>2</td>
<td>0.845</td>
<td>1893</td>
<td>7</td>
<td>-</td>
</tr>
<tr>
<td>NGC_6913</td>
<td>16</td>
<td>0.744</td>
<td>1148</td>
<td>&lt;13</td>
<td>M29</td>
</tr>
<tr>
<td>NGC_7243</td>
<td>10</td>
<td>0.220</td>
<td>808</td>
<td>114</td>
<td>-</td>
</tr>
</tbody>
</table>

Table - B1: The table of the cluster stars that we observed from McDonald Observatory, Fort Davis, TX using the 2.1m telescope

We computed the equivalent width (EW) of the H – Balmer lines for the stars from the flux absent in the lines and fit them with a Voigt profile. The EW includes both the Hydrogen (H) absorption line and the contributions of the metal lines. When H
absorption line is fit with a Voigt profile, the EW is almost completely the absorption from the H line, and no metal line contribution. Therefore, the ratio of the Voigt EW to the H line EW is less than 1 for cool stars, where the metal lines are stronger. Then they approach 1 as the H line grows. This is not only because the temperature makes the metal lines weaker, but because the hydrogen line becomes so enormous that the metal lines are only a very small percentage of the EW. Beyond the temperature of about 8750 K, the hydrogen line again begins to shrink, but now the star is so hot that the metal lines are gone and the ratio remains "1".

A plot of EW-ratio for the cluster stars against the EW Voigt value with the isochrones is shown below:

![Fig B1: Plot of EW-ratio for the cluster stars against the EW Voigt value with the isochrones](image)
The reason that we have much fewer stars in the above plot than what we originally observed is because a large percentage of our observed stars aren't actually members of the clusters. After we took the data, a new database (MWGSSC – Milky Way Global Survey of Star Clusters II – Kharchenko et al. 2013) was released that used a combination of proper motions and color-magnitude diagrams, including 2MASS colors, and many of our highly confident cluster members (which we chose from proper motions only) actually turned out to not be cluster members.

The blue dotted line separates the region where Helium line is detected. Stars falling above this line are very hot stars (Teff > 11,000 K). The intention was to use these stars as calibrators, for instance of the isochrones. However, most of the cluster stars are hotter than 9000 K and the Canis Major sample is mostly cooler stars. So it didn't work too well.

Fig B2: A plot of BP stars and the various types of stars from the Miles spectral library
If we look at the above plot that shows the Canis Major BP stars and various types of stars from the Miles spectral library, we can clearly see that most of the BP stars lie along the trend where the Miles main sequence stars are.

If the cluster stars that we observed were main sequence stars we could perform the similar test using the EW of the H – Balmer lines and B – V values as we did for the BP stars explained in Appendix A. In that case, if we did not have a significant difference between the reddening values obtained from using our suggested method and the known reddening values for the cluster stars then we can say our technique is significantly correct and we can use this in computing the reddening values for the BP stars.

At McDonald Observatory we got mostly the brighter stars that were hotter stars and most of them turn out to not be main sequence stars. We also lost a lot of our samples because of the result of the MWGSSC survey. We really need the main sequence stars to solve the reddening problem and one way to get the spectra for the main sequence stars is to get some time on a multi-fiber object spectrograph.
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- Sharoz Rafiul Islam, Mirza; Wilhelm, R. J., “Spectral results for the blue plume stars in Canis Major Overdensity” American Astronomical Society, AAS Meeting #223, #355.08