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Jaká kritéria jsou vhodná ke klasifikaci hvězd horní části hlavní posloupnosti?

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Jakub Faktor

Vedoucí práce: doc. Mgr. Ernst Paunzen, Dr. **Brno 2016**

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Abstrakt

Klasifikace hvězd je v současnosti jednou z nejvíce prominentních disciplín astrofyziky, jež se rapidně vyvíjí již od publikace Yerkes klasifikačního systému v roce 1943. Kritéria klasifikace byla za uplynulé roky prodiskutována, některá byla zanechána, jiná zůstávají až dodnes. Tato práce zvažuje kritéria aplikovatelná při klasifikaci spekter hvězd hlavní posloupnosti typů O, B a A ve viditelné části spektra až po spektrální čáru Hβ, zejména pak dle flexibilních MK standardů Pickles nízkého rozlišení. Za tím účelem byl v programovacím jazyku C# vyvinut software SpecOp. Dnes jsou počítače schopny modelovat hvězdné atmosféry, ale hvězdná spektra jsou stále často vyhodnocováno manuálně, pomocí tabulkových editorů nebo jednoduchých programů. SpecOp je plně rozvinutá aplikace s grafickým rozhraním, obsahující několik knihoven hvězdných spekter, která je schopná vykreslit a zpracovat hvězdné spektrum a určit ekvivalentní šířky důležitých spektrálních čar.

Abstract

The stellar classification is presently one of the most prominent disciplines of astrophysics, that has been developing rapidly, ever since the Yerkes classification system was published in 1943. The criteria of the classification were broadly debated over the years, some were left behind and some remain to this day. This thesis ponders on the various criteria applicable in the classification of spectra of main sequence O, B and A type stars in the visible region up to the spectral line Hβ, particularly according to the flexible low resolution Pickles MK standards. For this purpose, the software SpecOp was developed in the C# programming language. Today, computers are able to model stellar atmospheres, yet stellar spectra are still often evaluated manually, using table editors, or simple programs. The SpecOp is a fully fledged application with graphical interface, containing multiple stellar spectra libraries, that is capable of plotting and processing stellar spectra and determining equivalent widths of important spectral lines.

Místo tohoto listu vložte kopii oficiálního (podepsaného) zadání práce.

Poděkování

Poděkování patří především vedoucímu práce, doc. Mgr. Ernstu Paunzenovi, Dr., za vedení, ochotu a čas, který mi po dobu zpracování práce věnoval.

Prohlášení

Prohlašuji, že jsem svoji bakalářskou práci vypracoval samostatně s využitím informačních zdrojů, které jsou v práci citovány.

Brno 16. května 2016 . . . .

Jakub Faktor

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Introduction

Spectroscopy today is a highly developed method of studying electromagnetic radiation and stands as method bringing a vast amount of information about distant objects to us. Product of a spectroscopic measurement is called a spectrum and it is essentially a function of distribution of radiation respective to wavelength. Analysis of measured spectra is one of the main tasks of astrophysics. Let it be stars, nebulae or galaxies that we observe, spectroscopy always gives us some useful information, that furthers our understanding of physics of the observed object. Spectroscopic methods are of great importance especially when it comes to classifying stars according to their spectra. Stellar spectra can tell us much not only about the physical nature of observed stars, like temperature, surface gravity and chemical composition, but also give us a crucial piece of information needed to identify larger celestial systems and structures, for instance determining the radial velocity or the rate of rotation. However, to gain this information, we need to classify the star most of the time, in order to determine what its spectra look like without the various deformations by phenomena like Doppler's effect or cosmological redshift. That is achieved through the use of stellar classification, which is mostly based on classifying stars according to ratios of relative intensities of specific spectral lines. In the present, the majority of stars are classified according to the Yerkes system. The classification is done by determining both the spectral and luminosity class, effectively describing the star's position in the Hertzsprung-Russell diagram. Observable spectra are different for each spectral type and so the ratios of intensities of lines used for classification are type-specific. However, precision of any classification using this method is dependent on the specific lines used and while our current theory of stellar evolution can be used to choose correct criteria, we must be mindful of anomalous objects that would flaw our classification results.

The aim of this thesis is an attempt to compile a set of stellar classification criteria used for classification of hot stars and classify a set of selected stars according to it, using software developed for this purpose. Studied spectra were selected from A library of stellar spectra, George H. Jacoby, Deidre A. Hunter and Carol A. Christian, 1984, [1] while while the versatile spectra of A Stellar Spectral Flux Library: 1150 – 25000 Å, Pickles A. J., 1998, [2] were used as standard spectra. The software SpecOp was developed in the C# programming language and was made to be lightweight, easily-accessible and controllable application capable of semi-automatic to automatic MK system stellar classification for the Windows NT operation systems. This thesis is only considering criteria applicable to main sequence stars of types O, B and A in the visible region up to the Hβ spectral line.

Chapter 1

Spectroscopy

1.1 Historical background

When Isaac Newton first used a prism to disperse a beam of light in the 17th century, he could not have known the importance of his discovery. As he began studying the newly found spectrum, as he called it, spectroscopy was born. Many great minds of physics studied the phenomenon throughout the history in attempt to explain its physical nature and improve the methods used in spectroscopic measurements. Josef von Fraunhofer used concepts of Thomas Young's theory of interference to increase the precision of spectroscopic measurements by replacing the dispersing prism by a diffraction grating. Later on, he made systematic observations of the Sun, where he, in the year 1814, observed dark lines that bear his name to this day. However, what he was unaware of, was the origin of these lines. The fact that there is some relation between the measured spectrum and the chemical composition of the observed material was discovered independently by Léon Foucault and Anders Angström. This relation was first quantitatively explained in 1859 by Gustav Kirchoff and Robert Bunsen. They postulated that every element and compound has its own characteristic spectrum and observed flames produced by many elements, including Lithium, Potassium, Strontium, Calcium and Barium. [3] In the second half of the 19th century studies of black body radiation by Gustav Kirchhoff and Ludwig Boltzmann advanced our understanding of spectra. Max Planck showed that the energy transmission is not continuous, but travels in the form of energy quanta and in the year 1900 he achieved a breakthrough by developing the equation that we know today as Planck's Law. The path to modern spectroscopy was paved by the works of Niels Bohr, mainly by his 1913 Theory of atomic hydrogen, which revolutionized not only spectroscopy, but all of physics. [4]

1.2 Electromagnetic radiation

Before proceeding to description of spectra and their origin, some characteristics of electromagnetic radiation should be noted. Electromagnetic radiation is, according to Maxwell theory of electromagnetic field, a transverse wave propagating through vacuum at the speed of light. The frequency ν [Hz] and wavelength λ [nm] of electromagnetic radiation are tied by the relation (1.1):

$$c = \nu\lambda, \tag{1.1}$$

where c is the speed of light. Electromagnetic radiation is formed by photons, elementary boson particles (thus possessing an integer spin, 0 specifically for photons) that have a rest mass equal to zero. Max Planck showed that energy E [J] of an energy quantum, a photon, is given by the equation (1.2):

$$E = h\nu = \frac{hc}{\lambda}, \tag{1.2}$$

where h is Planck's constant, ν [Hz] is frequency and λ [nm] is wavelength of the radiation. In the physics of stars, electromagnetic radiation has the major role of the medium of energy, carrying the energy from the inner regions to the outer regions and even to the star's surroundings. Furthermore, it is the source of radiation pressure, which in turn has a major role in the physics of hot stars. [5, 6]

1.3 Black body radiation

All objects emit light and the hotter and more luminous they are, the more they radiate. To quantitatively describe radiation that is being emitted, the physical model of the absolutely black body is used. The model consists of several constraints: it is defined as a non-reflective body that's surface is uniformly held at constant temperature. Absolutely black body is a physical approximation applicable in many scientific fields. It is here, that thermodynamics meet atomic physics. For astrophysics it holds major importance, as absolutely black body is a rather accurate approximation for stars and other astrophysical objects of interest, even though stars are not completely non-reflective, nor are they in thermal equilibrium with their surroundings. It was Kirchhoff's 1859 work that showed that the black body radiation only depends on the temperature of the absolutely black body's surface and not on the properties of the surface itself. [5] The black body radiation spectrum is in the form of continuum, consisted in a continuous interval of wavelengths, with spectral distribution changing according to temperature. Generally, black body radiates more and in higher frequencies (differently put, in shorter wavelengths), the greater its temperature is. Quantitatively, this is described by Planck's Law (1.3):

$$B_\nu(\nu,T) = 2\pi \frac{\nu^2}{c^2} \frac{h\nu}{exp(h\nu/kT)-1}, \tag{1.3}$$

where B_ν is monochromatic intensity in specific radiation frequency and temperature, ν is frequency, T is temperature of the black body, h is Planck's constant and k is Boltzmann's constant. [5] This equation leads to the characteristic shape of spectral distribution of black body radiation (fig. 1.1), that is shown below:

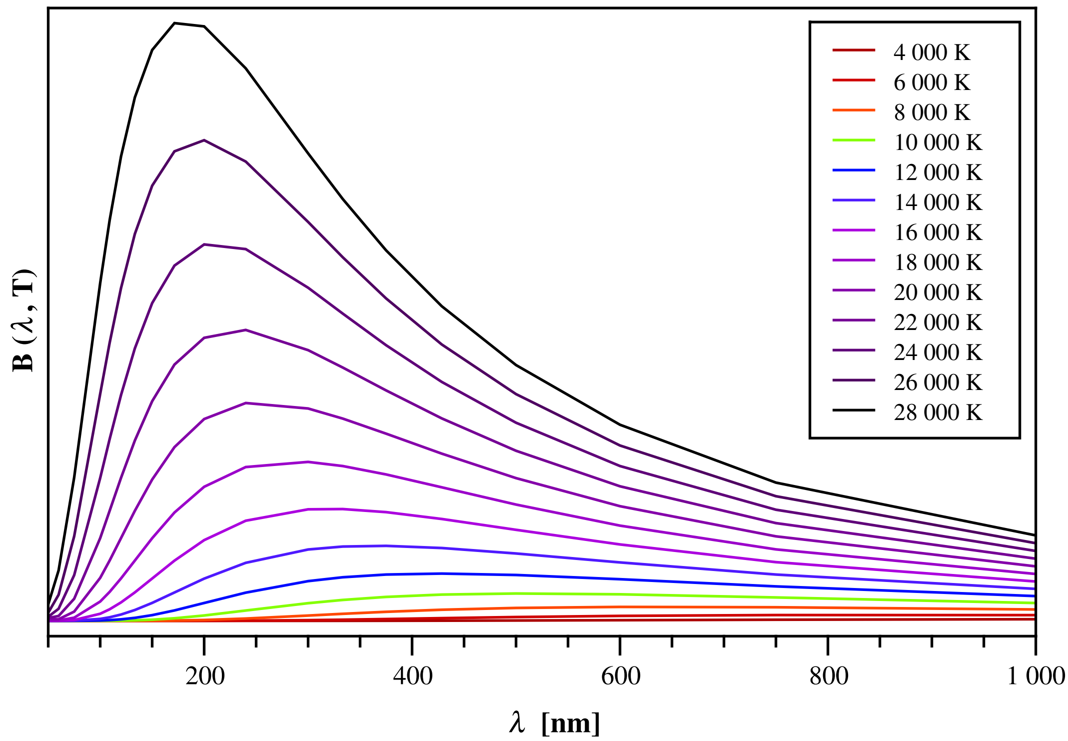

Figure 1.1: Spectral distribution of a black body at several values of temperature T [K].

From this graphical representation, it's easy to notice the fact that the wavelength of the maximum of the spectral distribution is decreasing with increasing temperature, as higher energy radiation is being radiated. Wien's displacement law is an example of an approximation that makes use of this fact. [7] Luminous energy that is radiated by the black body is tied to the black body's effective temperature T_{eff} [K] by the Stefan-Boltzmann's Law (1.4). [5]

$$\Theta_e = \sigma S T_{eff}^4, \tag{1.4}$$

where σ is Stefan-Boltzmann constant and S is the radiating surface area of the black body. If we consider that we intend to approximate stars as spherical black bodies we arrive at the equation (1.5), that describes luminosity L [$\mathrm{Js^{-1}}$] of a star with radius R [m] and effective temperature T_{eff} [K]:

$$L = \sigma T_{eff}^4 4\pi R^2. \tag{1.5}$$

1.4 Photon absorption and emission

Most often absorption / emission of a photon is product of change of momentum of a charged particle. The lighter the charged particle is, the easier it is to change its momentum and the easier a photon is emitted. In practical astrophysics, radiation that is product of changes in the movement state of electrons is studied nearly exclusively. [5] In the case of electrons bound in atoms, the electron's movement state is represented by its quantized energy levels. Fundamental characteristic of photons is the fact that their energy is always exactly equal to the energy difference of the emitting electron's initial and final energy states. In the case of absorption of photon by electron bound in an atom, the final state of the atom is the more energetic of the two energy levels and the term excitation is used. In the majority of cases the atom only remains in an excited state for a relatively short amount of time, before an opposite effect of de-excitation takes place. As stated, only a photon with energy equal to the energy difference between initial and final energy states of the electron can be emitted. This, of course, holds true for the inverse scenario, where a photon can only be absorbed by an electron, if the electron's energy levels allow it. This leads to the distinctive shape of atomic spectra, that is formed by a collection of spectral emission and absorption lines. If, on the other hand, the energy of interacting charged particle is not limited to discreet distribution, for instance a situation in which a free electron is changing its momentum after entering an electrostatic field, energy of the photon that is emitted will depend on the initial energy of the charged particle. Thus, if the studied matter allows its charged particles' energy continuous intervals of possible energy values, for example in the state of plasma, the spectrum is in the form of continuum. [5, 8]

1.4.1 Spectral lines

Both absorption and emission spectral lines are product of bound-bound transitions between two states of an atom. The nomenclature is apparent, if an atom absorbs a photon, the spectra contain absorption lines, while if the atom emits a photon, we observe emission lines. Sets of hydrogen and helium emission lines, which are product of de-excitation of an atom to a specific energy level, are ordered in series named after their discoverers. Hydrogen spectral series are, in the order of ascending principal quantum number, Lyman series (n = 1) in the ultraviolet region of the spectra, very important Balmer series (n = 2) in the visible region, and Paschen series (n = 3), Brackett (n = 4), Pfund series (n = 5) and Humphreys series (n = 6) in the infrared region. Further series are not named. As goes for helium line series, which are not quite as famous as hydrogen line series, the named series are Fowler series (n = 3) and Pickering series (n = 4), which, originally Edward Pickering wrongly attributed to some unknown state of hydrogen. Lines belonging to the series are then labeled with Greek characters, for instance the first Lyman line, which is product of n = 2 to n = 1 transition of hydrogen atom, is named Lα, the first Balmer line (n = 3 to n = 2) is Hα and so on. Knowledge of specific energy levels of studied spectral line enables us to calculate its wavelength. For single electron atoms, including hydrogen and singly ionized helium atoms, we can use equation (1.6):

$$h\nu = \Delta E = Z^2 E_1 \left(\frac{1}{n_i^2} - \frac{1}{n_j^2} \right) \rightarrow \frac{1}{\lambda} = -\frac{E_1}{hc} Z^2 \left(\frac{1}{n_i^2} - \frac{1}{n_j^2} \right) = Z^2 R \left(\frac{1}{n_i^2} - \frac{1}{n_j^2} \right), \quad (1.6)$$

where h is Planck's constant, ν is frequency of the photon, E_1 is the energy of the ground state, c is the speed of light, n_i and n_j are the principal quantum numbers of the transitioning states and R is the Rydberg's constant. The wavelength difference for the neighbouring lines in a series is decreasing with the higher principal quantum number n_i, until they are zero in the case $n_i \rightarrow$ infinity. Here the spectral line series turn into a continuum and the wavelength at which this continuum begins is defined as the series limit. [5, 6]

The fact that spectral lines form due to precisely defined transitions would seem to suggest that the energy and thus the wavelength of spectral line photons is a specific value, however, due to quantum mechanical principles, both the energy and the wavelength of these photons are an interval of values, we can say that the lines are not perfectly sharp and that they posses certain profiles. In order to evaluate individual line profiles, usually equivalent widths of the spectral lines are used. [5] Equivalent width W_λ [Å] at wavelength λ [Å] is a measure of a photometric strength of the studied line and is formaly defined with the equation (1.7):

$$W_\lambda \equiv \int_0^\infty (1 - r_\lambda) d\lambda, \tag{1.7}$$

where r_λ is residual flux, which we gain after normalizing the flux values of the spectrogram. The equivalent width can be interpreted as the length of a rectangle possessing height equal to one, that spans the same area as the profile of the spectral line as shown in the illustration (1.2). [9]

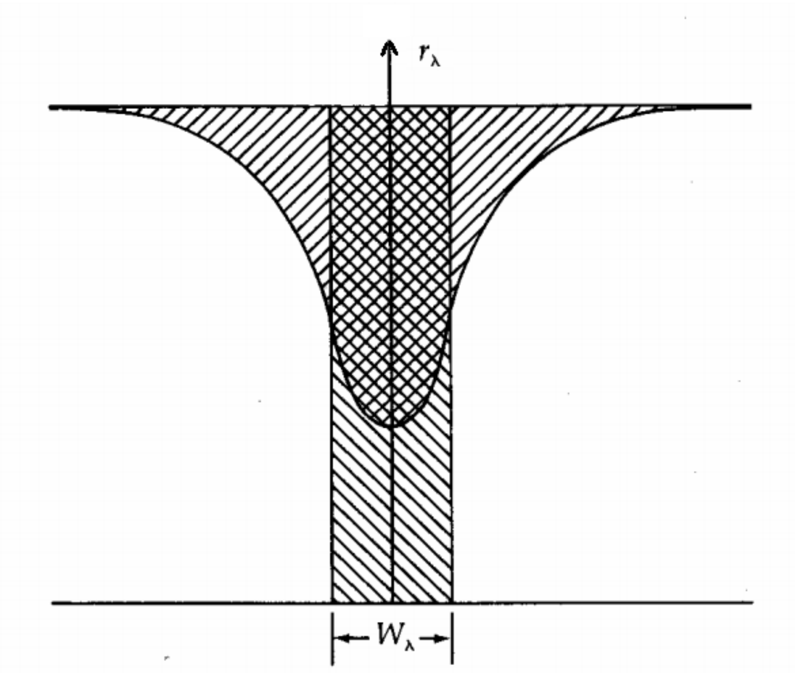

Figure 1.2: Equivalent width W_λ of an absorption spectral line. [9]

1.4.2 Continuum

Both the series limit and following continuum are labeled according to the related series, hence Lyman limit, Lyman continuum, Balmer limit, Balmer continuum etc. Apart from the continua related to the series, continua that are product of multitude of other phenomena can be observed as well. These phenomena include bound-free transitions and free-free transitions. In bound-free transitions the electron that was being held by the nucleus is emitted with some kinetic energy, while also emitting a photon accordingly. This is what we call ionization and the reverse, where an ion catches an electron, is called recombination. In free-free transitions, an electron moving through space near an ion emits a photon, yet still has high enough kinetic energy to escape. [5]

1.4.3 Transitions in stars

The stellar spectra are not so much a product of the chemical composition of the observed star, but rather are determined by the star's effective temperature. Transitions in stars happen in cycles, progressing in a rapid succession of events. Neutral atoms get ionized, ionized atoms recombine, atoms get excited / de-excited. However, if the temperature does not change, the ratios of atoms in specific states are invariant in time. This state of distribution is called statistical equilibrium. For N_m and N_n particles in the states m and n in statistical equilibrium, the ratios of N_m to N_n can be calculated using Boltzmann's equation (1.8):

$$\frac{N_m}{N_n} = \frac{g_m}{g_n} exp\left(-\frac{E_m - E_n}{kT}\right), \tag{1.8}$$

where g_m and g_n are the state statistical weights, given by the level of degeneracy of the given level, E_m and E_n are energies at said states, k is Boltzmann's constant and T is what we call excitation temperature. The ratio between number of atoms that are *i+1* times ionized and number of atoms i times ionized, again in the state of statistical equilibrium, can be expressed using Saha's equation (1.9):

$$\frac{N_{i+1}}{N_i} = \frac{2}{N_e}\frac{Z_{i+1}}{Z_i}\left(\frac{2\pi m_e kT}{h^2}\right)^{3/2} exp\left(\frac{E_i}{kT}\right), \tag{1.9}$$

where Z_i is partition function of given level of ionization, N_e is the concentration of free electrons, E_i is the energy of the specific ionization, h is Planck's constant and T in this equation is called ionization temperature. The ratios of number of atoms in specific states is important for stellar classification due to the fact, that the spectral distribution, or more precisely, the ratios of relative intensities, that are being studied, is depending on the number of atoms of each present element in each state. [5]

1.5 Properties of a spectrograph

Spectrographs, the scientific instruments used to measure observed light's spectral distribution, the spectra, exist in many different designs, providing measurements of varying qualities. Apart from the major dividing characteristic of type of dispersing element used, spectrographic systems possess further properties, most importantly dispersion D and resolving power R resulting in spectral resolution $\Delta\lambda$.

1.5.1 Dispersion

The sole goal of a spectrographic instrument is to isolate and measure monochromatic components of the studied radiation, in other words, its spectra, described by radiation wavelength and luminosity. The physical effect of splitting a light beam into a multitude of monochromatic beams of different angles relative to the original beam's direction is called dispersion, or sometimes chromatic dispersion. The greater these angles are, the more spread the measured spectral lines are, and the greater is the dispersion D [$^\circ m^{-1}$] of the measuring system. It is defined (1.10) in the following way:

$$D = \frac{\Delta\theta}{\Delta\lambda}, \tag{1.10}$$

where $\Delta\theta$ [°] is the angular spread of the neighbouring spectral lines and $\Delta\lambda$ [nm] is the difference of the wavelengths of said lines. [6] In astrophysics, different values of dispersion are used when observing spectral lines of different strengths. If the measured spectra were obtained in low dispersion, only the strong spectral lines will be present in the spectrogram. Higher values of dispersion will yield results with weaker lines observable alongside the strong ones. This thesis studies the classification criteria applicable at different values of used dispersion: 125 Åmm^{-1} [10], 75 Åmm^{-1} [11] and 40 Åmm^{-1} [12], which are discussed further in the chapter 2.

1.5.2 Dispersing element

The dispersing element is a crucial part of any spectrograph, as means to separate the component wavelengths of the studied radiation. Generally, spectrographs can be divided into two groups: instruments using either a prism or a diffraction grating as a dispersing element. In modern times, the former has been largely superseded by the latter, mainly due to the higher possible resolving power of the diffraction grating. The two types of dispersers physically work in different ways. While prisms work based on refraction of light, diffraction gratings make use of interference of light. This fundamental difference gives the two dispersing counterparts several important traits. [8]

1.5.3 Spectral resolution

Quality of a spectrogram is given by the measure of possibility of resolving detailed spectral features in the measured spectra, hence the name of this property – spectral resolution $\Delta\lambda$. This characteristic is strongly tied to the resolving power R of the used instrument. The relation of spectral resolution and resolving power of a spectrograph can be described by the equation (1.11):

$$R = \frac{\lambda}{\Delta\lambda}, \tag{1.11}$$

where spectral lines of wavelengths λ and $(\lambda + \Delta\lambda)$ are the closest possibly resolved lines of the spectrogram. [6] It was earlier stated that Josef von Fraunhofer increased the precision of spectroscopic measurements, when he replaced the dispersing prism by a diffraction grate. To articulate this in less vague terms, this substitution allowed designing of spectrographs with greater resolving power, resulting in spectrograms of higher spectral resolution. [8]

1.5.4 Prism

As prisms as dispersing elements are based on refraction of light, the main characteristic of a dispersing prism is its refractive index $n(\lambda)$, which is related to the speed of light c inside the prismatic medium and dependent on the incoming wavelength λ. Snell's law, derived from Fermat's principle, describes the angle of refraction of light (and in turn dispersion) on the refracting interface between two media (1.12):

$$n_1(\lambda)\sin\alpha = n_2(\lambda)\sin\beta, \tag{1.12}$$

where $n_1(\lambda)$ and $n_2(\lambda)$ are the refractive indices of the two media, α is the angle between the direction of incoming light and the direction of a normal vector of the surface of the interface and β is the angle between the direction of the beam of the refracted radiation and the direction of the normal vector of the interface surface. As for spectral resolution, the equation (1.13) [8] applicable in this case is:

$$R = b\frac{dn}{d\lambda}, \tag{1.13}$$

where R is the resolving power, b is a constant and $dn/d\lambda$ is the rate of change of refractive index with respect to wavelength. It is apparent that the component $dn/d\lambda$ is directly proportional to the resolving power, but it should also be noted that this factor depends on the material of the disperser. This fact is fundamental to the limit of resolving power of any spectrograph that makes use of a dispersing prism, as we lack materials that would accommodate our spectrographic needs. [8]

1.5.5 Diffraction grating

Another approach in submitting radiation to chromatic dispersion is the use of the fore-mentioned diffraction grating. It consists of a series of either closely spaced parallel slits or grooves ruled on a hard glassy or metallic material. As the term suggests, diffraction gratings make use of diffraction of light. Diffraction is the phenomena, that occurs when radiation passes through a slit of a size similar to the radiation's wavelength, as a consequence of the wavelike character of radiation and the principles of interference. General equation (1.14) of diffraction on a grating states:

$$m\lambda = d(\sin i + \sin\theta), \tag{1.14}$$

where i [°] and θ [°] are the angles of incidence and reflection, d [mm^{-1}] is the groove spacing that represents the proximity of individual grooves, λ [nm] is the wavelength of radiation and m is an integer quantity, that is called the order of diffraction. Diffraction gratings as dispersing elements do not disperse the incident light into merely one spectral projection as do prisms. Instead, multitude of spectral projections, hence orders of diffraction, is created by the interaction of the radiation after passing through the grating. [8] The equation of dispersion then becomes (1.15):

$$D = \frac{m}{d\cos\theta}, \tag{1.15}$$

which shows the inverse dependence of dispersion D on groove's spacing θ. [6] Instead of dispersion D, the inverse measure called plate factor P [Åmm^{-1}] is more often used. [13] It is therefore described by the equation (1.16):

$$P = \frac{dcos\theta}{m}. \tag{1.16}$$

The equation of spectral resolution can be similarily rewritten as (1.17):

$$R = Nm, \tag{1.17}$$

where N is the number of diffraction grating's grooves. [6] It is apparent, that for a set order of diffraction m, the spectral resolution R increases with the number of grooves N of a spectrograph linearly.

Chapter 2

Stellar Classification

2.1 History of stellar classification

The first attempt at creating a system of stellar classification was made by Angelo Secchi in 1868. He published a catalogue consisting of 4 000 stellar spectra at low dispersion and proposed a system of four classes denoted by Roman numerals: I – white stars with hydrogen lines, II – yellow stars with metallic lines, III – orange stars with absorption lines and IV – red stars with absorption lines sharp at their high wavelength side and blurred at their short wavelength side. In 1890 Secchi's classification was superseded by the Harvard classification scheme, that was created by Edward Pickering and Williamina Fleming and later refined by Antonia Maury. The refined class system became O-B-A-F-G-K-M, which we today know to be the progression order of descending temperature, the original alphabetical progression was based on the strength of hydrogen lines. In the years 1890 to 1924, the Harvard classification underwent further development and had a major role in the creation of the Henry Draper catalogue – the famous HD catalogue, consisting of about half a million stars with identified spectral types in even a more refined system where each class was divided into ten subclasses temperature-wise, for instance F0 being the hottest spectral class F star and F9 being the coolest spectral class F star. [5]

The problem of the Harvard classification lies in its single parameter criterion – temperature. The Hertzsprung-Russell diagram is a perfect display of the reason why this is a problem. It is apparent, that the distribution of stars in the HR diagram as a whole cannot be viewed upon as a function of only temperature, because the mapping of temperature to luminosity would violate the definition of function, where every element of the function's codomain is the image of at most one element of its domain. In the year 1943 a new system, named the Yerkes spectral classification, but also called MKK, bearing the names of its creators William Wilson Morgan, Philip C. Keenan, and Edith Kellman, was introduced as the first two-parametric model. The new system used both temperature (spectral classification) and luminosity (luminosity classification) in order to classify the star and quickly started gaining popularity. In 1953 Morgan and Keenan published a revised version of MKK, now named MK system, which is still used to this day.

2.2 Morgan-Keenan system

The Morgan-Keenan system is a phenomenology of spectral lines, blends, and bands, based on general progression of colour index and luminosity. It is defined by an array of standard stars, located on the two-dimensional spectral type vs luminosity-class diagram. These standard reference points are not depending on values of any specific line intensities or ratios of intensities; they have come to be defined by the appearance of the totality of lines, blends and bands in the ordinary photographic region. [14]

In this way, the authors themselves retrospectively defined their creation in their 1973 MK system edition. One of the most crucial thoughts of the original introduction of the MK system is consisted in the definition, that is, the emphasis on study of stars and their spectra in the relation to the many different groups of stars, separated by differently looking spectra, while using as little theoretical assumptions as possible. It is obvious, that as there are very many stars, the best course of action is to attempt to find these characteristic stellar specimen, that can be used as a solid general representation of the group and in turn use them to classify the less typical ones. [15] Hence, in the MK system, standard stars are used to classify the stars of unknown type. MK's precursor, the MKK system, published by William W. Morgan, Philip C. Keenan and Edith Kellman in the year 1943, was a vast improvement of the Harvard classification system used at the time, as it brought the second parameter of luminosity classification. After being published, the two-dimensional classification system underwent more improvements and ten years later, re-named as Morgan-Keenan system, gained worldwide acceptance. It did not mean that the MK system has reached its final form, on the contrary the system thrived, as many upstanding individuals of the scientific community sought to refine it even further.

As it was mentioned earlier, MK system is defined at specific value of dispersion 125 $\text{Å}\text{mm}^{-1}$. This value was originally discussed in the publication of MKK – An Atlas of Stellar Spectra, 1943.

The dispersion used (125 $\text{Å}\text{mm}^{-1}$ at $H\gamma$) is near the lower limit for the determination of spectral types and luminosities of high accuracy. The stars of types F5-M can be classified with fair accuracy on slit spectra of lower dispersion, but there is probably a definite decrease in precision if the dispersion is reduced much below 150 $\text{Å}\text{mm}^{-1}$. The lowest dispersion capable of giving high accuracy for objective-prism spectra is higher; the limit is probably near 100 Åmm^{-1}. The minimum dispersion with which an entirely successful two-dimensional classification on objective-prism plates can be made is probably near 140 Åmm^{-1}. This value was arrived at from a study of several plates of exquisite quality taken by Dr. J. Gallo, director of the Astronomical Observatory at Tacubaya, Mexico. [16]

2.3 Spectral classification

The first dimension of classification according to the MK system is essentially a present day, refined, form of the Harvard classification, defining a set of spectral classes with effective temperature decreasing in the order of progression: O-B-A-F-G-K-M, or more precisely with the supplementary classes added throughout the modern era of astrophysics (fig. 2.1):

$$
\begin{array}{cccccccccc}
 & & & & & & \mathrm{R} & \mathrm{N} & & \\
\mathrm{W} & \mathrm{O} & \mathrm{B} & \mathrm{A} & \mathrm{F} & \mathrm{G} & \mathrm{K} & \mathrm{M} & \mathrm{L} & \mathrm{T} \\
 & & & & & & & \mathrm{S} & &
\end{array}
$$

Figure 2.1: Hardvard classification with supplementary classes.

As for the early types, O and B type stars have strong helium lines and A type stellar spectra have the strongest hydrogen spectral lines. W type stars are the hot, massive and very luminuous Wolf-Rayet stars, originally massive O type stars that lost a portion of their mass through strong stellar winds and their spectra exhibit similar lines to O type spectra, but some lines of strongly ionized carbon, oxygen and nitrogen are also observable. Spectral lines of metals gain in strength towards the late types, as do spectral lines of molecules. R and N type stars are sometimes called C type or carbon stars, cold stars with strong carbon lines in their spectra, S type stars are similar to the giant M type stars, except for the fact that their spectra contain strong zirconium oxide instead of titanium oxide lines, T type stars are brown dwarfs with strong methane spectral lines. The overview of spectral types, their intervals of temperatures and observable spectral lines is listed in the table (2.1) below [5, 17]:

Spectral type	Temperature [K]	Spectral lines
W	> 25 000	He II, He I, H I, O III-VI, N III-V, C III, C IV, Si IV
O	28 000 - 50 000	He II, He I, H I, O III, N III, C III, Si IV
B	10 000 - 28 000	He I, H I, C II, N II, Fe III, Mg III
A	7 300 - 10 000	H I, ionized metals
F	6 000 - 7 300	H I, C II , Ti II, Fe II
G	5 300 - 6 000	Ca II, neutral metals, simple molecules
K	3 800 - 5 300	Ca I, neutral metals, molecules
M	2 500 - 3 800	Ca I, molecular bands of TiO
L	1 300 - 2 100	TiO, VO, FeH, CrH, H_2O
T	600 - 1 300	CH_4, Na I
R	3 655 - 5 350	CH, CN, C_2
N	2 940 - 5 030	CH, CN, C_2, SiC_2, C_3
S	2 400 - 3 500	Ca I, ZrO

Table 2.1: Table listing spectral types, their effective temperatures and associated strong lines.

An even better overview of the dependency of strength of spectral lines of specific elements on spectral type can be gained from the graphical interpretation (fig. 2.2) of some of the strong lines [17]:

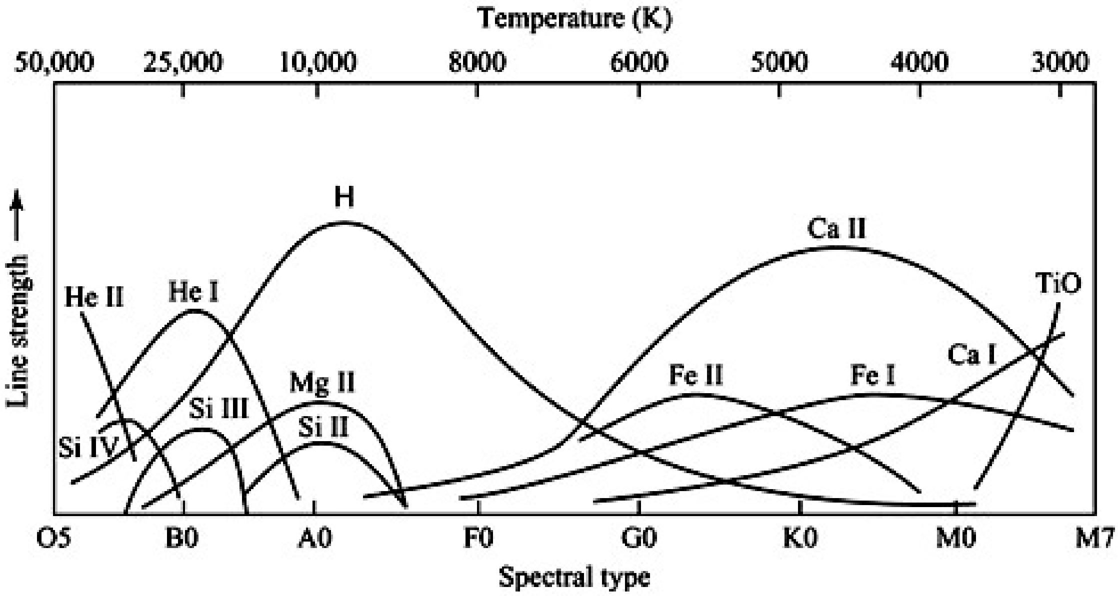

Figure 2.2: Illustration of the presence of strong spectral lines in the temperature progression. Figure taken from: `https://cmswork.nau.edu/CEFNS/Labs/Meteorite/About/Glossary-Ss/`.

As it can be seen, stars with higher effective temperature radiate at spectral lines of higher ionization order and elements with higher ionization energies. The major thought behind the general approach of using ratios of intensities of specific spectral lines can be seen in this illustration. Combinations of ratios of spectral lines are virtually always unique to their temperatures, even if the ratios themselves are not unique or if they reach similar values. Some of the general characteristics of the stellar spectra are contained here as well, for instance strong helium lines of class O and B stars, strong Hydrogen lines of the A type stellar spectra, strong ionized Iron and Calcium lines of type F and G stars and so on. A more detailed description of the characteristic spectra of the different stellar types is given in the chapter 4.

While ratios of intensities at the wavelengths of selected spectral lines should be in principle usable for the classification, or its estimate, at the very least, the most common, scientific way of evaluating the differences between the tested spectra and the standard spectra is using the ratios *W'* of equivalent widths *W* of selected lines instead. Such ratio is traditionally formally described, in the likeness of the following ratio (2.1), proposed by Conti and Alschuler, [18] with the spectral line identifiers included [18, 19]:

$$W' = \frac{W(4471\ \ He\ \ I)}{W(4541\ \ He\ \ II)}. \tag{2.1}$$

2.4 Luminosity classification

The luminosity classification, describing the total radiation output of a star, brings the second parameter of luminosity into stellar classification. The method is based on the fact that the measured stellar spectra not only give us information about the effective temperature, but also information about the surface gravitational acceleration. Spectra of stars with lower surface gravitational acceleration possess sharp and thin spectral lines, while spectra of those with larger surface gravitational acceleration have broad spectral lines, caused by frequent particle collisions. [5] This is due to the the gas pressure of the photosphere – that is a function of surface gravitational acceleration – and we call this effect pressure broadening.

The star's surface gravitational acceleration is tied to the mass of the star *M* and radius *R* by the equation (2.2). [20]

$$g = \kappa \frac{M}{R^2}, \tag{2.2}$$

where constant κ is the gravitational constant ($\kappa = 6.6742 \cdot 10^{-11}$ m^{3}kg^{-1}s^{-2}). In practice, however, the surface gravitational acceleration is more typically used in its logarithmic form (2.3): [21]

$$\log g = \log M - 2\log R + 4.437. \tag{2.3}$$

The MK system classes use eight luminosity classes in total: Ia – Luminous supergiants, Ib – Supergiants, II – Bright giants, III – Giants, IV – Subgiants, V – Main sequence stars, also called dwarfs, VI – Subdwarfs and VII – White dwarfs. Often, only classes I-V are used for practical reasons. The stars of each luminosity class typically have similar values of the logarithm of surface gravitational acceleration $\log g$: for the class I stars the $\log g \approx -0.5$, for the class II stars $\log g \approx 0.5$, for the class III stars $\log g \approx 1.5$, for the class IV stars $\log g \approx 3$ and finally, for class V stars, the point of the interest of this thesis, the $\log g \approx 4.5$. [22]

In the subchapter 1.3, it was shown how we can approximate a star as a black body using Stefan-Boltzmann's law. However, according to the equation, the star's luminosity is depending on the fourth power of its effective temperature and the radiating surface area S, which only contains the second power of its radius R. What this implies is that the luminosity is less influenced by the radius in hot stars and more influenced in cool stars. This fact also means that luminosities of stars of particular luminosity classes of hot stars differ less than those of cool stars. [7]

Chapter 3

General Characteristics of hot main sequence stars

3.1 Hot and cool star distinction

Earlier theories of stellar evolution supposed that stars evolutionarily are born as hot stars and progress by gradual cooling to cool stars. We know today that this is not correct, but the evolutionary series retained the labels that were used. And so, hot stars are sometimes called early, while cool stars are called late. The two groups were separated at the temperature of the Sun (5 780 K). Today we prefer to divide the stars into groups according to their physical characteristics, that we can observe (thus mostly at the outer layers) or quantify. These include the mass, luminosity, radius of the layer, from which the observed radiation is originating, the surface gravity and most importantly the effective temperature of the area of origin of observed radiation. With our current knowledge of stellar physics, it can be stated that the effective temperature is the factor that has the major role in both state of stellar atmospheres and inner regions of the star, that can not be observed directly. The fact that hotter stars also differ from the cool ones in the existence of local magnetic fields and active regions on the surface enables us to find a separation point between hot stars with calm atmospheres and cool atmospheres that possess these features. The role of effective temperature is even more determining in stars, where the energy flux rising from the inner regions is carried prevalently either by radiation or by convection. The former is true for the hot stars, whereas the latter is true for the cool stars. It turns out that the separation point fitting this criterion lies at the effective temperature of spectral type F2, which is approximately 7 000 K. Most often the group of hot stars is considered to consist of spectral types O, B and A, while F, G, K and M are considered cool stars. [20]

3.2 Hot main sequence stars

Hot stars are a diverse group of stars, that are typically very large and bright. Hot stars are rare due to their relatively short times spent on the main sequence. Even so, they form a significant portion of all of the observed stars, since their greater luminosity enables us to detect them from greater distances. The largest subgroup of hot stars is made up of hot main sequence stars. To attain a better perception of what these are, let us observe the famous Hertzprung-Russell (fig. 3.1).

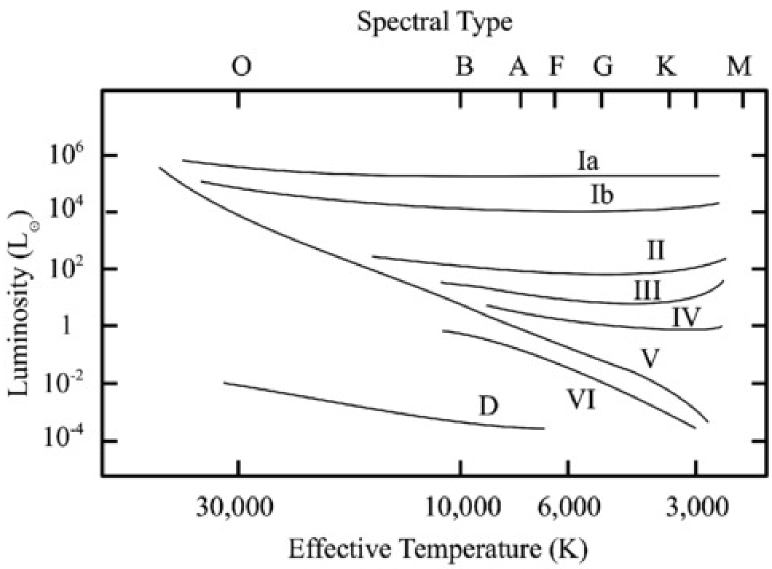

Figure 3.1: HR diagram showing the eight luminosity classes. Figure taken from: `https://ase.tufts.edu/cosmos/print_images.asp?id=49`.

This particular HR diagram was constructed using data from Schaller *et al.*, 1992 [23] and shows the evolution of stars of various initial masses M_0 over their lifetimes. Hot stars are situated at the left side of the HR diagram. If considering positions of spectral types O, B and A in the HR diagram, most are situated in the area of stars of the main sequence population I.

Main sequence stars are defined as stars that radiate exclusively at the expense of energy produced in the thermonuclear fusion. Among the fusion reactions in the hot main sequence stars dominates the CNO cycle, where carbon, nitrogen and oxygen act as catalyzers in the reaction of fusion of hydrogen into helium. The dominant type of thermonuclear reaction ongoing in the stellar core is dependent on the chemical composition of the core, but mainly on the core temperature. In cool stars, the low temperature does not allow CNO cycle reactions to occur, instead, energy is produced by other fusion reactions, mainly the proton-proton chain, which is another distinction between the two groups. Characteristics of the main sequence stars are primarily given by their mass, which is nearly invariant in time for most stars, their rotation and chemical composition are secondary factors. [20]

One of the things main sequence stars have in common, is that they follow the mass-luminosity relation (3.1) between the star's luminosity L and the star's mass M:

$$L \propto M^{\eta}, \tag{3.1}$$

where the slope η is only changing slightly with time. Why such aproximation is possible is shown in the illustration (3.2) below, that is depicting the dependency between the star's mass and luminosity of a collection of main sequence stars.

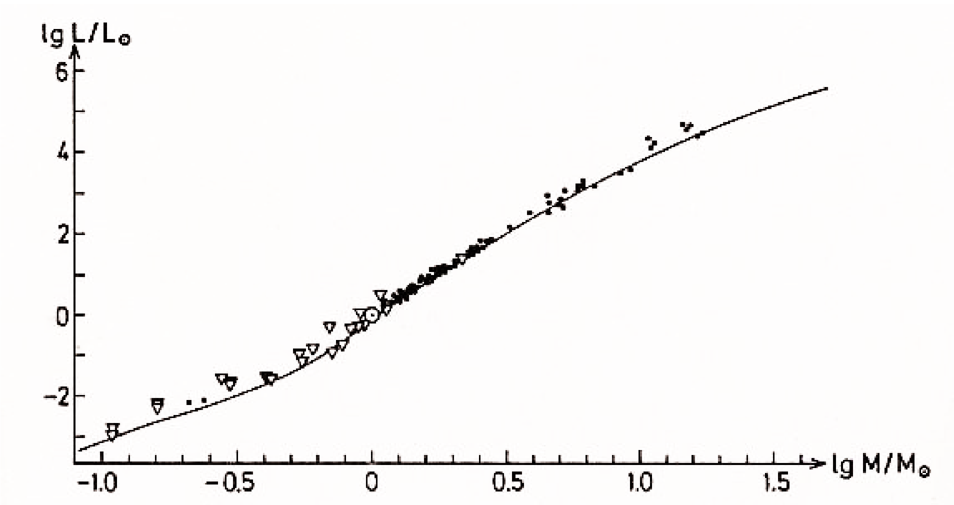

Figure 3.2: The mass-luminosity relation for main sequence stars. Figure taken from: `http://www2.astro.psu.edu/users/rbc/a534/lec18.pdf` [24].

For stars with masses between one and ten solar masses a solid general approximation lies at the value of $\eta \approx 3.88$ [24], for more precise approximation, the equations (3.2, 3.3) are sometimes used [25].

$$\frac{L}{L_{\odot}} = \left(\frac{M}{M_{\odot}}\right)^{4.0} (M > 0.43 M_{\odot}), \tag{3.2}$$

$$\frac{L}{L_{\odot}} = 0.23 \left(\frac{M}{M_{\odot}}\right)^{2.3} (M < 0.43 M_{\odot}). \tag{3.3}$$

In a similar way to the previous rule, the main sequence star's mass M and radius R are tied by the relation (3.4)

$$R \propto M^{\xi}, \tag{3.4}$$

however, the dependency in reality contains a significant change in slope around $M \approx M_{\odot}$, as can be seen in the illustration (3.3).

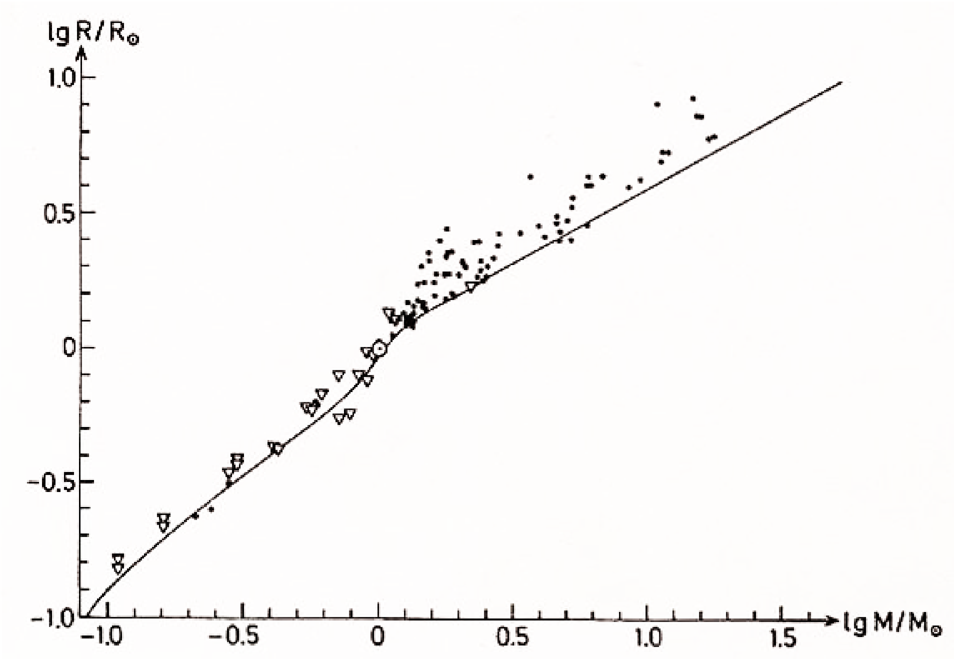

Figure 3.3: The mass-radius relation for main sequence stars. Figure taken from: `http://www2.astro.psu.edu/users/rbc/a534/lec18.pdf` [24].

Because of that, two different values of the slope ξ are usually used in practice: $\xi \approx 0.57$ for stars with $M < M_{\odot}$ and $\xi \approx 0.80$ for stars with $M > M_{\odot}$. This change of slope is caused by the fact, that the convective envelope, that is present below the star's photosphere, extends much deeper in the atmospheres of hot stars and the star's radius is then smaller, as the radiation pressure is lower due to radiation easier leaving the inner regions of the star. [24]

Chapter 4

O, B, A type main sequence stars and their classification

The following chapter more deeply describes the characteristics of the hot main sequence stars and their spectra and ponders on the applicable classification criteria. All of the shown spectrograms were obtained on the Gray / Miller spectrograph on the 0.8m telescope of the Dark Sky Observatory, using the 1 200 gmm^{-1} grating which gives a resolution of 1.8 Å. [26]

4.1 O type stars

O type stars with temperatures ranging from 28 000 K to 50 000 K [27] are massive (18 – 150 M_S), blue and hot. In fact, with the exception of Wolf-Rayet stars, O type stars are the hottest stars in the existence, which, of course, implies immense luminosities, especially if we consider their typically large radii ($>$ 6.6 R_S). Such losses of energy through radiation, of course, shorten the time the star exists as main sequence star, making O type stars very rare, as this time only amounts to millions of years. The wavelength at which O type stars radiate the most lies in the ultra-violet part of the spectrum. O type stars are often observed in OB associations, cosmologically compact areas in the giant molecular clouds, containing a large number of O type and B type stars, that share a common origin, but have gradually become gravitationally unbound, although they are still moving through space together. Stellar evolution of O type is happening at cosmologically rapid scales of millions of years.

Spectra of O type stars are defined as spectra simultaneously exhibiting helium lines in both neutral and ionized states. The intensity of the singly ionized helium line is maximal in the spectra of early O type stars, while the intensity of the neutral helium line grows towards its maximum in spectra of later O type stars [27]. This antagonistic behaviour is caused by the fact that the ratio of atoms in two said states is only a function of temperature as says the Saha's equation, while the total amount of helium atoms remains constant. This enables us to classify the star in terms of temperature and the MK system uses this principle for classification of O3-O9 stars [28]. Apart from helium, the spectra also contain weak lines of hydrogen. It can be noted that at the temperatures of O type stars, the majority of hydrogen in the photosphere is ionized and as such does not emit radiation. Furthermore, lines of ionized metals including Si IV, C III and O II. N III emission line can be present and if that is the case, while the spectrum also includes a He II emission line, the star is classified as Of. On the other hand, if the spectrum exhibits emission in Balmer and He lines, the star is classified as Oe. For the classification of later O type stars, namely O6-O9, ratio of C III 4649 / He II 4648 can be used for main sequence stars, if the ratio is less than one. Another possible criterion of classifying the O6-O9 stars is the ratio of Si IV 4089 / He I 4143 in high dispersion used by Conti and Alschuler, 1971. [18, 28]

As it was already stated, O type stars are characterized by weak hydrogen lines, lines of neutral helium He I and by lines of singly ionized helium He II, that can be observed in the figure (4.1).

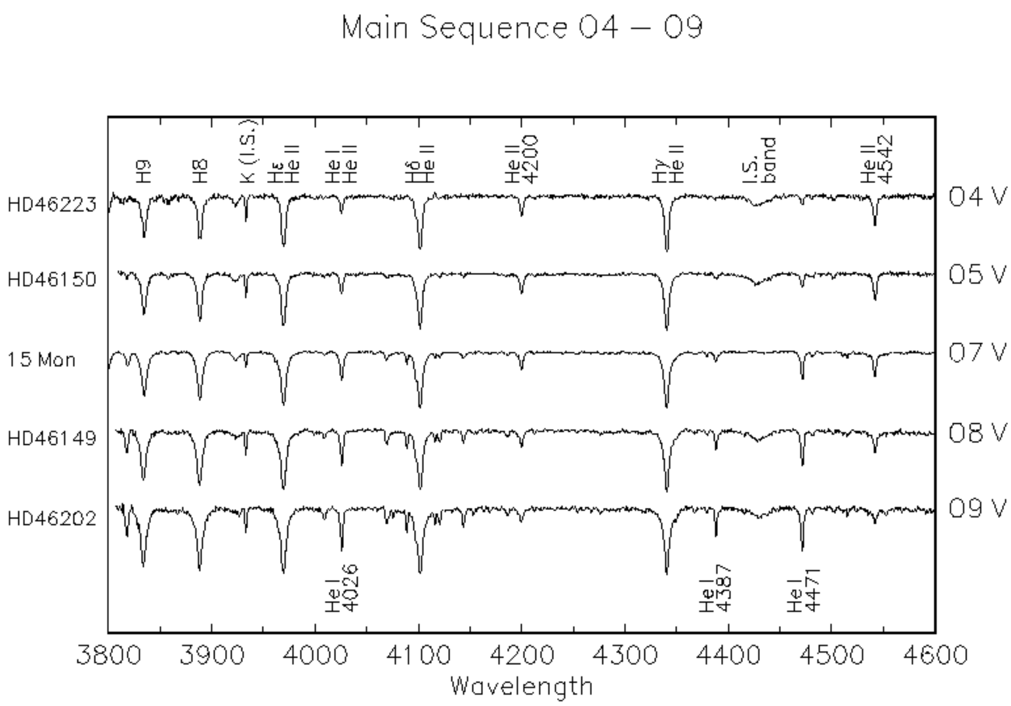

Figure 4.1: The O type main sequence stars. Figure taken from: `http://ned.ipac.caltech.edu/level5/Gray/frames.html`. [26]

The ratio He I 4471 to He II 4542 shows the antagonistic trend of neutral / singly ionized helium spectral line strengths. In several of the shown spectra two interstellar features can be observed. The Ca II K line, labeled "K (I.S.)", as well as the diffuse interstellar band, labeled "I.S. band" at a wavelength of about 4430 Å are both interstellar features attributed to interstellar gas. [26]

As goes for notable luminosity effects, at O9 the hydrogen lines become more sensitive to differences in luminosity than at earlier types, the spectra are shown in the figure (4.2) below. [26]

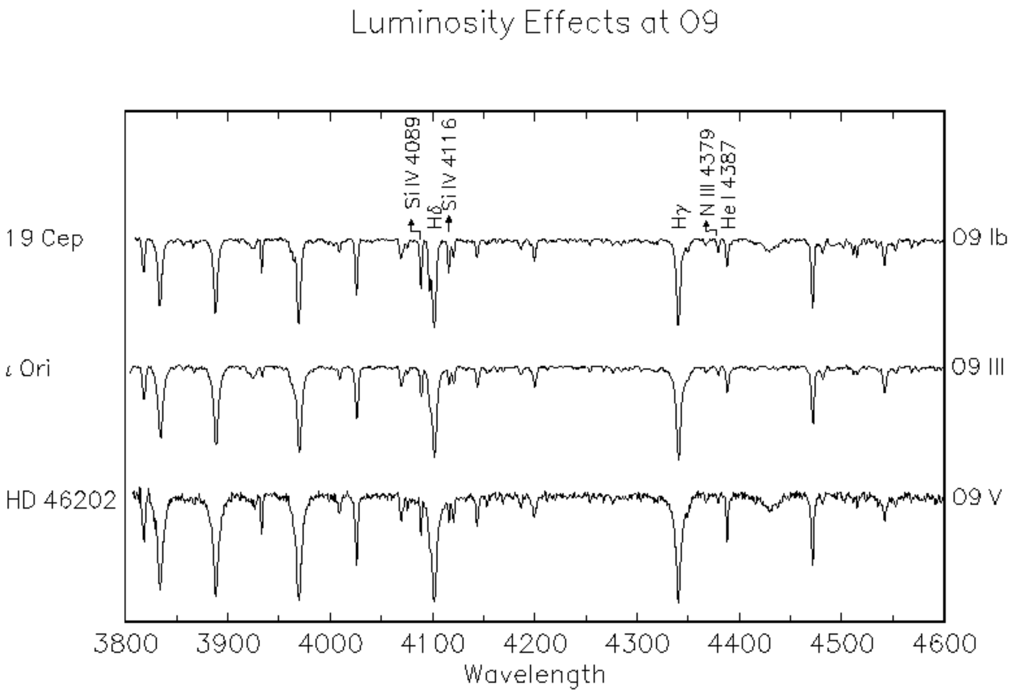

Figure 4.2: The luminosity sequence at O9. Figure taken from: `http://ned.ipac.caltech.edu/level5/Gray/frames.html`. [26]

After careful inspection, it can be noticed that, the ratio of Si IV 4089, which increases in strength as does luminosity, to Hδ, which becomes weaker with increasing luminosity, the ratio of Si IV 4116 to the neighbouring He I 4121 line, and the ratio of N III 4379 to He I 4387 pose as possible criteria for luminosity classification at spectral class O9. [26]

The helium lines of O type stars can be significantly influenced by metallicity effects: the smaller the value of metallicity, the stronger the ionized He absorbtion lines will be. The absorbtion lines of He I are affected more strongly than lines of He II. The He lines individually change according to effective temperature and luminosity class and, when used in certain ratios, may even shift the classification results to one subtype later. [29]

4.2 B type stars

B type stars are a group of blue-white hot stars with temperatures of 10 000 – 28 000 K, typically situated in OB associations. B type stars are typically slightly smaller (1.8 – 6.6 R_S) and less massive (2.9 – 18 M_S) than O type stars, their "lifetimes" spent as main sequence stars may, however, be many times longer.

The main difference between the spectra of O and B stars is in the presence / absence of lines of singly ionized helium. B type stellar spectra contain lines of helium exclusively in its neutral state. Also, in the spectra we can observe not only stronger lines of ionized hydrogen, but also lines of neutral hydrogen. In a similar way as ratios of ionized / neutral helium were used for classification of O type stars, ratios of H I / He I can be used for quantitative classification of B type stars. Some weaker lines in the wavelength range 3 600 – 4 800 Å can be used for the classification, but the number of observable elements drops towards the later B type stars and so all of the visible lines should be used in order to classify B5-A0 stars. The Yerkes system uses ratios of He I, Si II, III and IV and Mg II. [28, 30] This method of classification may, however, sometimes lead to flawed results, as the elements may behave in an abnormal way, for instance in helium stars and CNO objects. One of the approaches to this problem was only using He and Si lines, as suggested Nolan R. Walborn [31], stars with Si anomalies

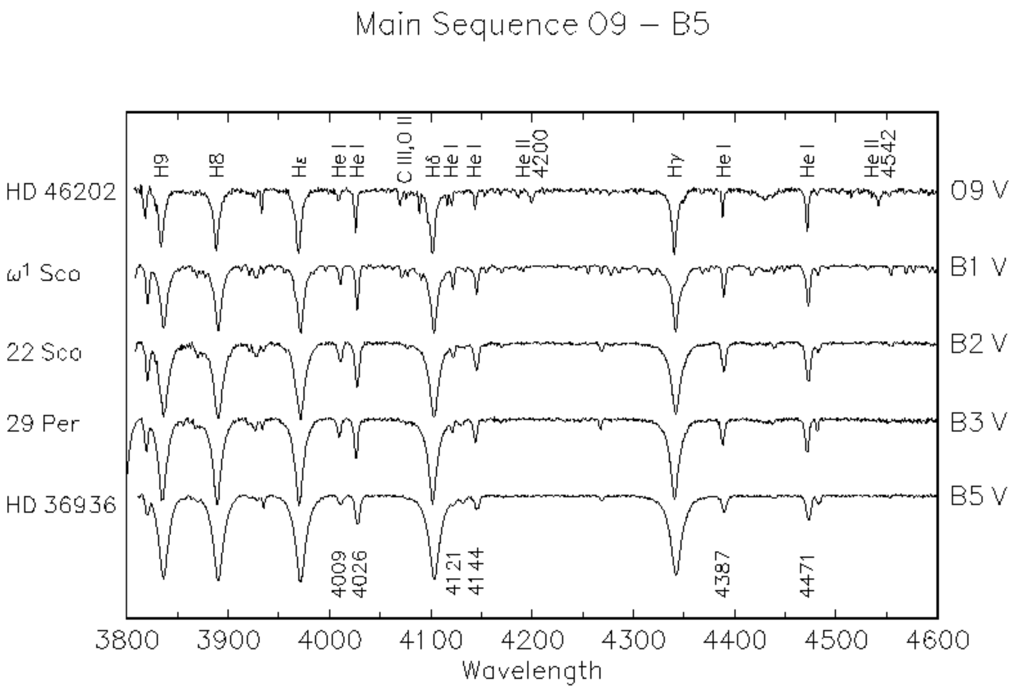

Figure 4.3: The main sequence from O9 to B5. Figure taken from: `http://ned.ipac.caltech.edu/level5/Gray/frames.html`. [26]

were not known yet at the time, but stars with He anomalies were known to exist and so the general approach became to use the rest of the ratios to confirm the classification result. [28] Figure (4.3) shows spectra of the early main sequence B type stars.

The intensity of the lines of He I reaches a maximum at approximately B2 and for cooler B type stars it weakens. A possible criterion of classification of spectral type lies in the ratio of He I 4471 / Mg II 4481 [26] and also N II 3995 / He I 4009 [30], although these spectral lines could be difficult to resolve on low resolution spectra.

Toward later types, the strength of helium lines decreases, until they disappear in spectra of this resolution at a spectral type of about A0. Even here, the validity of the ratio He I 4471 / Mg II 4481 holds. The threshold at the spectral class A0 is marked by the appearance of Ca II K-line, as can be seen in the figure (4.4). [26, 28]

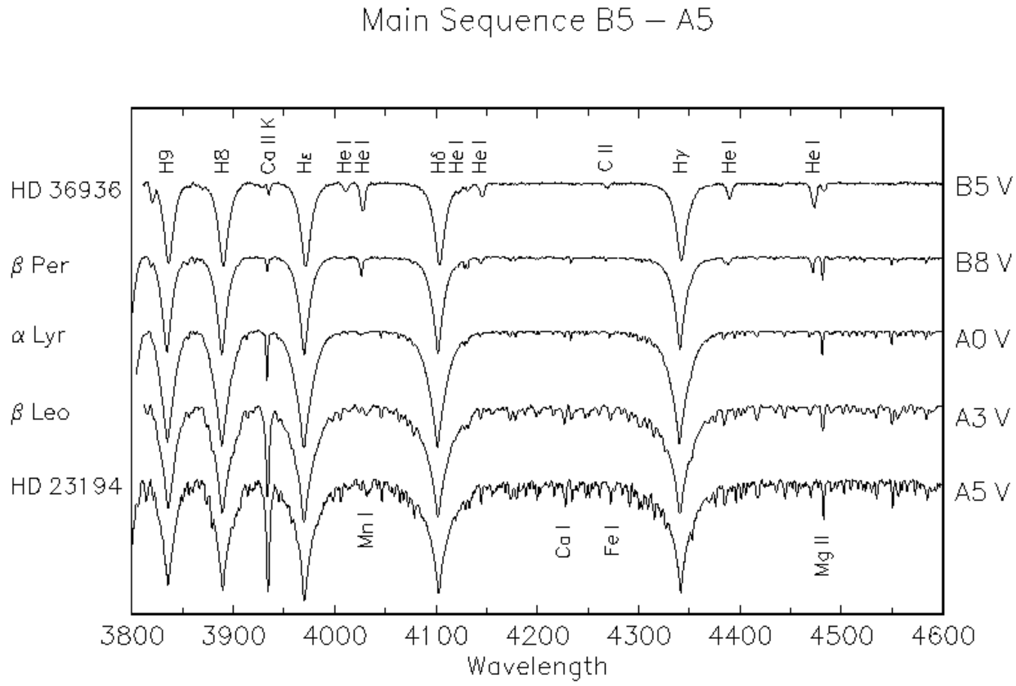

Figure 4.4: The main sequence from B5 to A5. Figure taken from: `http://ned.ipac.caltech.edu/level5/Gray/frames.html`. [26]

The luminosity classification is based on criteria that vary with the spectral type. The strength of the Balmer lines, which are weakening with increasing luminosity, holds information usable for luminosity classification, especially for the late B type stars. [28, 32]

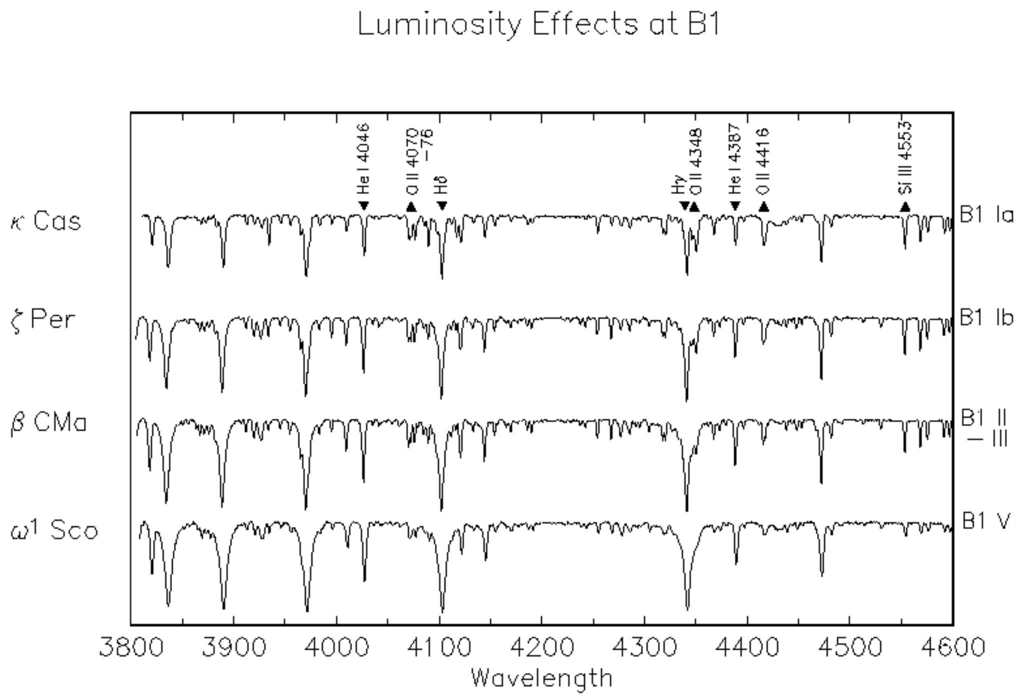

Figure 4.5: The luminosity sequence at B1. Figure taken from: `http://ned.ipac.caltech.edu/level5/Gray/frames.html`. [26]

The O II 4070, 4348 and 4416 lines are also strong for the early B type stars, as shown in the figure (4.5), and thus are considered to be usable in ratio with the Balmer lines. Notably, Si III 4553 line is used for B type stars as separating criteria between main sequence and giant luminosity classes. [26]

4.3 A type stars

A type stars are next in the temperature order, with temperatures ranging from 7 300 K to 10 000 K. A type stars are white, typically fast rotating stars, that are rather common, when compared to the rarity of stars of types O and B in the solar neighbourhood. A type stars, although the maximum of the spectral distribution of their radiation lies in the visible part of the spectrum, often also radiate strongly in the infrared part of the spectrum. This phenomena is attributed to a presence of dust in the system. [17]

A type stellar spectra exhibit strong absorption of the Balmer series, which is their prime characteristics as are strong lines of ionized metals. Hence, for classification of A type stars we can generally use ratios of relative intensities of lines of Hydrogen, Fe I, Fe II, Mg II, Ti II, Ca I and Sr II. In the spectra of early A type stars, neutral helium lines are present, although much weaker, when compared to B type stars. The neutral helium lines become no longer observable in the later A type stars. [5, 28] As it was mentioned in the B type section, the threshold at the spectral class A0 is marked by the appearance of Ca II K-line. Its strength increases further throughout the cooler A type stars, as can be seen in the figure (4.6). [26]

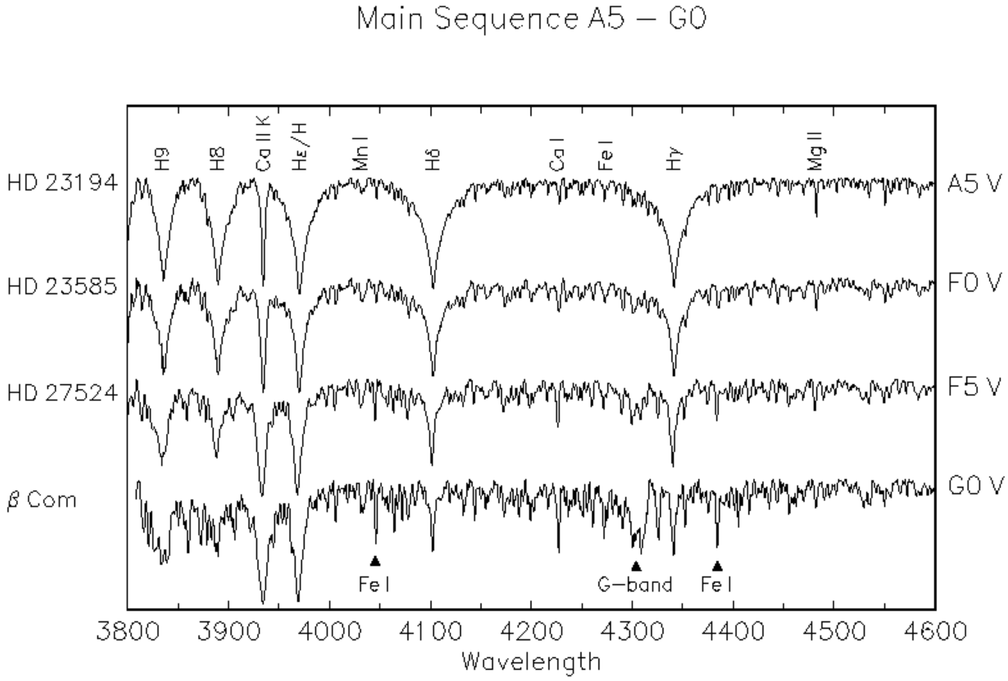

Figure 4.6: The main sequence from A5 to G0. Figure taken from: `http://ned.ipac.caltech.edu/level5/Gray/frames.html`. [26]

A note should be made at the Ca II K-line, as peculiar A type stars exist, in which these spectral lines are hardly observable. Similarly as the Ca II lines, the general metallic-line spectrum appears and strengthens at the A type stars. A spectral type classification criterion can be found in the ratios of lines such as Mn I 4030, Ca I 4227, and Fe I 4271. [26]

Chapter 5

SpecOp application

5.1 On the aim of the application

The SpecOp application was developed by me personally, for the purposes of this thesis, but also with the aim to make spectrum based stellar classification more accessible to both professional astrophysicists and astronomers and amateurs and also make it far less time consuming, while offering the convenience and user friendly graphical interface (fig. 5.2) the Microsoft Windows NT users are used to. The software was created in hopes that, when it is ready, it will be accepted by a university / scientific website and then will come to serve as a quality, actively updated, yet simple tool to users worldwide and its logo (fig. 5.1) shines proudly over the download links.

Figure 5.1: SpecOp logo.

5.2 On the current release SpecOp 0.7Beta

As the beta versioning suggests, the SpecOp is not at its final iteration yet as of this release. As this thesis was aimed at the classification of hot main sequence stars exclusively, so the rule sets and lists of spectral lines developed for the application were targeted at the O, B and A types.

The SpecOp application carries two stellar spectra libraries: A library of stellar spectra, George H. Jacoby, Deidre A. Hunter, and Carol A. Christian, 1984, [1] in the 3510 Å – 7427 Å region and An Atlas of Southern MK Standards From 5800 to 10200 Å, Anthony C. Danks, Michel Dennefeld [33]. Furthermore, a selection of the POLLUX synthetic spectra [34] and a set of universal MK standard star spectra published by Pickles, 1998 [2]. User can additionally load desired custom data in `.txt` files in ascii format.

The functions of SpecOp include loading, viewing, evaluating and plotting stellar spectra, all in the single frame standardized user environment (fig. 5.2). SpecOp is capable of estimating continuum of loaded spectra using the iterative peak striping method and normalize the data according to it. User can choose between two different matching methods, one for finding ratios of intensities at spectral line wavelengths, the other one for finding ratios of equivalent widths. For the determination of equivalent widths, the spectrum is scanned at wavelengths of the spectral lines of interest and found lines are measured with the use of block integration. After evaluating the ratios of found values, output tables of equivalent widths measured for each applicable spectral line as well as the ratios themselves are stored in a `.txt` file in ascii format.

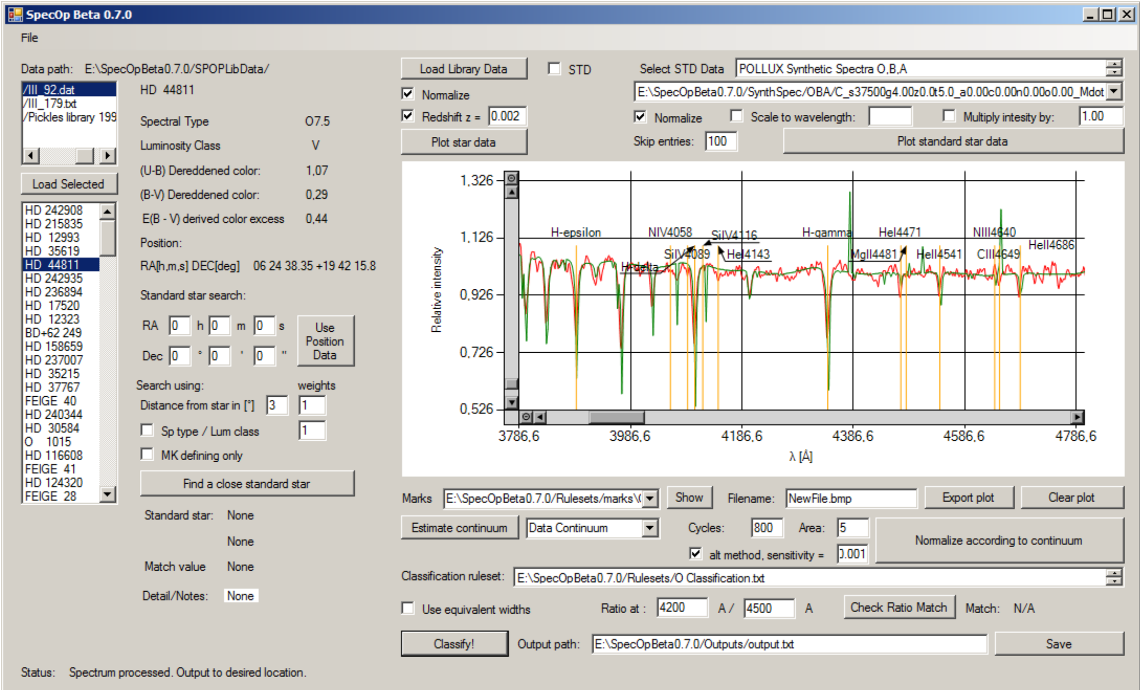

Figure 5.2: User interface of the SpecOp software.

The application was developed in the C# language using Visual Studio 2015. The digital size of the application is nearly negligible compared to the size of the packaged modules, that contain the library and catalog data, size of which may reach up 600 MB uncompressed.

For the addition of custom spectral data, additional loose `.txt` files, containing spectra wavelength and flux values, separated by a tab character, may be added in the ascii format to the `SpecOp\SPOPPhotData\SingleSpectra\` folder along with an appropriate line addition in the `ZZ_custom.txt` header file.

5.3 Peak stripping method

The application is using so called peak stripping method to estimate the continuum of a spectrum, which iterates over the values of wavelength and compares the values of flux at given wavelength with the mean flux value of the area, with the area interval being set by a parameter. For instance, the most simple and common version of the peak stripping method only uses the mean value $m_i = (y_{i-1} + y_{i+1})/2$ and then takes the greater of m_i and y_i. In this way, the spectrum is iterated trough many times, for this specific use of this method, 300 to 1000 cycles are made to gain an estimate of the continuum. [35] The illustration 5.3 shows the continuum estimates after (a) 300, (b) 500, (c) 800 iterations. The spectra in this thesis that were processed using the standard method were achieved using estimated continua generated after 500 cycles.

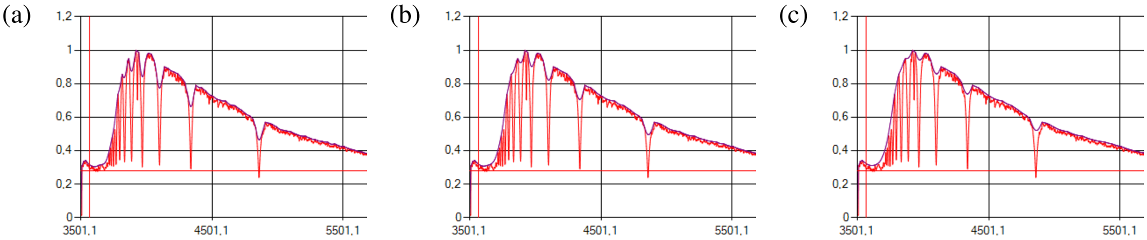

Figure 5.3: The use of the standard peak striping method after (a) 300, (b) 500 and (c) 800 iterations.

The standard version of the peak stripping method works well on spectra containing mostly absorption spectral lines, but its estimate may fade out any present emission spectral lines. Modifications of this method exist, that can be used to gain better estimates of continua of spectra containing both absorption and emission spectral lines, but the quality of the fit varies depending on the resolution of the spectra and parameters used. If the star is suspected of containing emission spectral lines in its spectra, that are essential to its classification, it would be reasonable to consider using such modification of the peak stripping method in addition to the standard one. In this thesis, apart from the standard one, a version of this method is used, where, in each cycle, values of y_i are replaced with m_i, if the formula $|y_i - m_i| < c$ is true, where c is a sensitivity parameter that enables us to adapt to the spectra of different properties. The modified peak stripping method that was used to process some of the spectra shown below was run with the parameters of 1000 cycles and sensitivity c = 0.0001, averaging over the area of nine values.

5.4 Future of the application

While developing the software, great care is given to the application architecture, so that more functionality, including, but not limited to, capability of classification of cooler stars, peculiar stars as well as stars of other Luminosity classes may be added by expanding code functions and parsers at pre-written list layers. In the future, various extensions of the applications will be released, including more specialized classification rules and standard libraries as well as new versions of the SpecOp software, that will bring new features and fix any possible bugs.

Results

Tested and standard spectra

The studied spectra, that were chosen from A library of stellar spectra, George H. Jacoby, Deidre A. Hunter, and Carol A. Christian, 1984, which extends from the wavelengths of 3510 Å to 7427 Å. The resolution of the spectra is $\sim$ 4.5 Å and the uncertainty of the photometric data is below 1 %. [1] As for the standard stars, the Pickles MK standards from A Stellar Spectral Flux Library: 1150 – 25000 A, Pickles A. J., 1998, were chosen, primarily for their flexibility. The resolution of the Pickles standard spectra is $\sim$ 500 Å. [2]

Important spectral lines

In order to make classification as precise as possible, the wavelengths of associated spectral lines were researched and corrected to a precision of .1 Å using tables of atomic spectra [36] and [37].

In the chapter 3, the use of spectral line C III 4649 for classification of late O type stars was mentioned. However, such spectral line is not present in [36] or [37]. The fact that two C III lines are close to the wavelength, C III 4650.25 and C III 4647.42, suggests the idea that the mentioned C III 4649 is actually a blend of these two lines. The decimal places of the wavelength associated for the use with the SpecOp software with the line were approximated to 4648.8 Å.

The fore-mentioned spectral line O II 4348 also seems to be this case, being a blend of O II 4347.43 and O II 4349.44. [37] The wavelength associated with the 4348 blend has been approximated as 4348.4. However the spectral line or rather the blend itself is very close to the much stronger Hγ line, making its use for classification very difficult, to the point it was excluded from the spectral lines system the SpecOp software uses. The blend is not very strong and is only considered as a classification criterion if it is an emission line, which may prove problematic if we only use the standard method. This fact suggests the use of the modified peak stripping method separately of the main classification to determine the presence of this line in emission.

Hδ and Hε lines in the studied spectra are very strong and might hold some classification value, for instance, for the classification of O stars the Hδ line would seem to be a valuable asset. Unfortunately, in spectrograms of MK dispersion, the Hδ spectral line unavoidably overlaps with a He II line, that we would probably like to use in ratio with, creating a blend. The Hδ + He II blend can still hold some information, but its interpretation could prove difficult, as the trend of increasing or decreasing spectral line strength with changing temperature becomes more complex. In the spectra of B and A type stars the He II line is not present and this problem therefore does not arise. Despite their limited use, both spectral lines are interesting features in the studied region and as such were added to the list of measured spectral lines.

Spectral lines used for each spectral type are listed with their associated wavelengths in the table (I).

O		B		A	
spectral line	λ [Å]	spectral line	λ [Å]	spectral line	λ [Å]
Hε	3889.0	Hε	3889.0	Hε	3889.0
N IV λ4058	4058.6	Ca II "H"	3933.7	Mn I λ4030-4	4032.0
Si IV λ4089	4088.9	N II λ3995	3995.0	Hδ	4101.7
Hδ	4101.7	He I λ4009	4009.3	Sr II λ4215	4215.5
Si IV λ4116	4116.1	He I λ4026	4026.2	Ca I λ4226	4226.7
He I λ4121	4120.8	O II λ4070-6	4073.0	Fe I λ4271	4271.5
He I λ4143	4143.8	Si IV λ4089	4088.9	Ti I λ4300	4300.1
Hγ	4340.5	Hδ	4101.7	Hγ	4340.5
He I λ4471	4471.7	Si IV + He I λ4121	4120.8	Fe I + Fe II λ4383-5	4384.0
Mg II λ4481	4481.2	Si II λ4128-30	4130.9	Fe II + Ti II λ4417	4416.7
He II λ4541	4541.6	He I λ4144	4143.8	Mg II λ4481	4481.2
N III λ4640	4640.0	Hγ	4340.5	Hβ	4861.3
C III λ4649	4648.8	He I λ4387	4387.9	Hα	6562.8
He II λ4686	4685.8	O II λ4416	4417.0		
		He I λ4471	4471.7		
		Mg II λ4481	4481.2		
		Si III λ4552	4552.6		
		Hβ	4861.3		
		Hα	6562.8		

Table I: Table of spectral lines measured for each type. [36, 37]

Compiled classification criteria

Of the researched spectral and luminosity classification criteria, the accepted and used ones are listed in the table (II).

	O		B		A	
	Spectral	Luminosity	Spectral	Luminosity	Spectral	Luminosity
Main criteria	N IV / N III He II / He I Si IV / Si III	C III / He II Si IV / He I	Si IV / Si III Mg II / H	O II / He I profiles of balmer lines	Ca I / H Fe I / H	profiles of balmer lines Fe I / H
0		C III λ4649 blend / He II λ4686	Si III λ4552 / Si IV λ4089 Mg II λ4481 / He I λ4471 Hδ / He I λ4471	Si IV + He I λ4121 / He I λ4144 He I λ4026 / O II λ4070-6 He I λ4387 / O II λ4416 Si IV λ4089 / He I λ4009	Ca I 4226 / Hβ, Hγ Mn I 4030-4 / Hβ, Hγ Fe I λ4271 / Hβ, Hγ Ti I λ4300 / Hβ, Hγ	Hδ / Hγ Hγ / Hβ Hβ / Hα
1		C III λ4649 blend / He II λ4686	Si III λ4552 / Si IV λ4089 Mg II λ4481 / He I λ4471 Hδ / He I λ4471	He I λ4026 / O II λ4070,76 He I λ4387 / O II λ4416 N II λ3995 / He I λ4009	Ca I λ4226 / Hβ, Hγ Mn I λ4030-4 / Hβ, Hγ Fe I λ4271 / Hβ, Hγ Ti I λ4300 / Hβ, Hγ	Fe II + Ti II λ4417 / Mg II λ4481
2	He I λ4471 / He II λ4541 N IV λ4058 / N III λ4640	C III λ4649 blend / He II λ4686 He I λ4471 / He II λ4541	Si III λ4552 / Si IV λ4089 Si II λ4128-30 / Si IV + He I λ4121 Mg II λ4481 / He I λ4471 Hδ / He I λ4471	He I λ4026 / O II λ4070,76 He I λ4387 / O II λ4416 Si IV + He I λ4121 / He I λ4144 N II λ3995 / He I λ4009	Ca I λ4226 / Hβ, Hγ Mn I λ4030-4 / Hβ, Hγ Fe I λ4271 / Hβ, Hγ Ti I λ4300 / Hβ, Hγ	Fe I + Fe II λ4383-5 / Mg II λ4481
3	He I λ4471 / He II λ4541 N IV λ4058 / N III λ4640	C III λ4649 blend / He II λ4686 He I λ4471 / He II λ4541	Si II λ4128-30 / Si IV + He I λ4121 Mg II λ4481 / He I λ4471 Hδ / He I λ4471	Hδ / Hγ Hγ / Hβ Hβ / Hα	Ca I λ4226 / Hβ, Hγ Mn I λ4030-4 / Hβ, Hγ Fe I λ4271 / Hβ, Hγ Ti I λ4300 / Hβ, Hγ	Fe II + Ti II λ4417 / Mg II λ4481
4	He I λ4471 / He II λ4541 N IV λ4058 / N III λ4640	C III λ4649 blend / He II λ4686 He I λ4471 / He II λ4541	Mg II λ4481 / He I λ4471 Hδ / He I λ4471	Hδ / Hγ Hγ / Hβ Hβ / Hα	Ca I λ4226 / Hβ, Hγ Mn I λ4030-4 / Hβ, Hγ Fe I λ4271 / Hβ, Hγ Ti I λ4300 / Hβ, Hγ	Fe II + Ti II λ4417 / Mg II λ4481
5	He I λ4471 / He II λ4541 Hδ + He II λ4100 / He II λ4541	C III λ4649 blend / He II λ4686	Si II λ4128-30 / He I λ4144 Mg II λ4481 / He I λ4471 Hδ / He I λ4471	Mg II λ4481 / He I λ4471 Hδ / Hγ Hγ / Hβ Hβ / Hα	Ca I λ4226 / Hβ, Hγ Mn I λ4030-4 / Hβ, Hγ Fe I λ4271 / Hβ, Hγ Ti I λ4300 / Hβ, Hγ	Fe II + Ti II λ4417 / Mg II λ4481
6	He I λ4471 / He II λ4541 Hδ + He II λ4100 / He II λ4541	C III λ4649 blend / He II λ4686 Si IV λ4089 / He I λ4143	Si II λ4128-30 / He I λ4144 Mg II λ4481 / He I λ4471	Hδ / Hγ Hγ / Hβ Hβ / Hα	Ca I λ4226 / Hβ, Hγ Mn I λ4030-4 / Hβ, Hγ Fe I λ4271 / Hβ, Hγ Ti I λ4300 / Hβ, Hγ	Sr II λ4215 / Ca I λ4226
7	He I λ4471 / He II λ4541 Hδ + He II λ4100 / He II λ4541	C III λ4649 blend / He II λ4686 Si IV λ4089 / He I λ4143	Si II λ4128-30 / He I λ4144 Mg II λ4481 / He I λ4471	Hδ / Hγ Hγ / Hβ Hβ / Hα	Ca I λ4226 / Hβ, Hγ Mn I λ4030-4 / Hβ, Hγ Fe I λ4271 / Hβ, Hγ Ti I λ4300 / Hβ, Hγ	Sr II λ4215 / Ca I λ4226
8	He I λ4471 / He II λ4541	C III λ4649 blend / He II λ4686 Si IV λ4089 / He I λ4143	Si II λ4128-30 / He I λ4144 Mg II λ4481 / He I λ4471	Hδ / Hγ Hγ / Hβ Hβ / Hα	Ca I λ4226 / Hβ, Hγ Mn I λ4030-4 / Hβ, Hγ Fe I λ4271 / Hβ, Hγ Ti I λ4300 / Hβ, Hγ	Sr II λ4215 / Ca I λ4226
9	He I λ4471 / He II λ4541 Mg II λ4481 / He I λ4471	C III λ4649 blend / He II λ4686 Si IV λ4089 / He I λ4143 Si IV λ4089 / Hδ Si IV λ4116 / He I λ4121	Mg II λ4481 / He I λ4471	Hδ / Hγ Hγ / Hβ Hβ / Hα	Ca I λ4226 / Hβ, Hγ Mn I λ4030-4 / Hβ, Hγ Fe I λ4271 / Hβ, Hγ Ti I λ4300 / Hβ, Hγ	Sr II λ4215 / Ca I λ4226

Table II: Compiled criteria used for spectral and luminosity classification of O, B, A subtypes. [10, 11, 12, 26, 27, 28, 30, 31, 32]

Star: HD 242908
Standard star: MK O5V

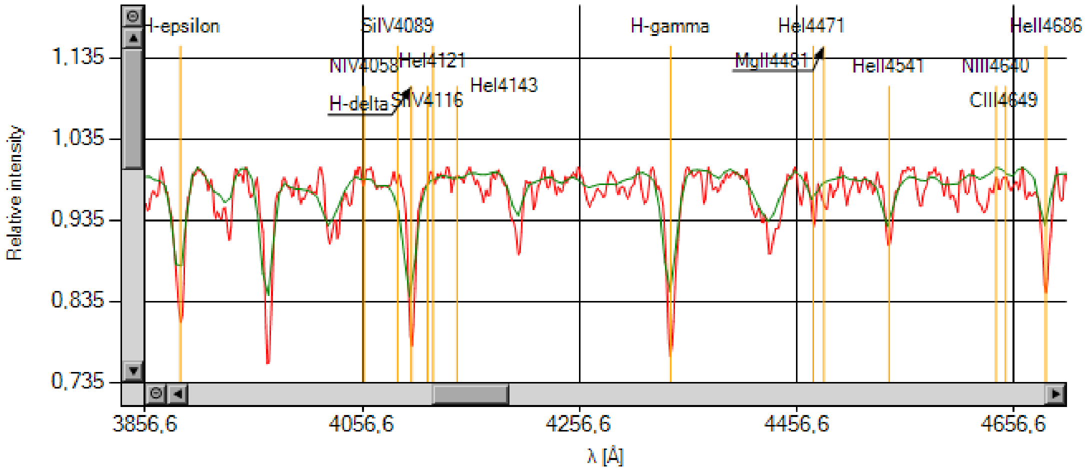

Figure I: Spectra of O5V stars: HD 242908 (red line) and MK O5V (green line).

spectral line	W_{STAR} [Å]	$W_{STANDARD}$ [Å]
N IV λ4058	0.073	0.222
Si IV λ4089	0.067	not resolved
Si IV λ4116	0.035	not resolved
He I λ4121	0.069	0.172
He I λ4143	0.045	0.187
He I λ4471	0.272	0.452
He II λ4541	0.592	0.852
Mg II λ4481	0.188	not resolved
N III λ4640	0.067	0.045
Hε	1.927	1.774
Hδ	1.738	2.417
Hγ	2.254	2.291
C III λ4649	0.061	0.107
He II λ4686	1.318	0.786

(a) Table of important spectral lines and their equivalent widths measured in the spectra of HD 242908 and MK O5V.

criterion	W'_{STAR} [Å]	$W'_{STANDARD}$ [Å]	$\frac{W'_{STAR}}{W'_{STANDARD}}$
N IV λ4058 / N III λ4640	1.078	4.906	0.220
He I λ4471 / He II λ4541	0.459	0.530	0.866
Mg II λ4481 / He I λ4471	0.691	–	–
C III λ4649 / He II λ4686	0.046	0.136	0.338
Si IV λ4089 / He I λ4143	1.473	–	–
Si IV λ4089 / Hγ	0.030	–	–
Si IV λ4116 / He I λ4121	0.506	–	–
Hδ / He II λ4541	3.804	2.686	1.416

(b) Table of spectral line ratio criteria, their associated values and the ratio between the star's and the standard's criterion value.

Table III: The criteria for classification of HD 242908.

Star: HD 242908
Standard star: MK O5V

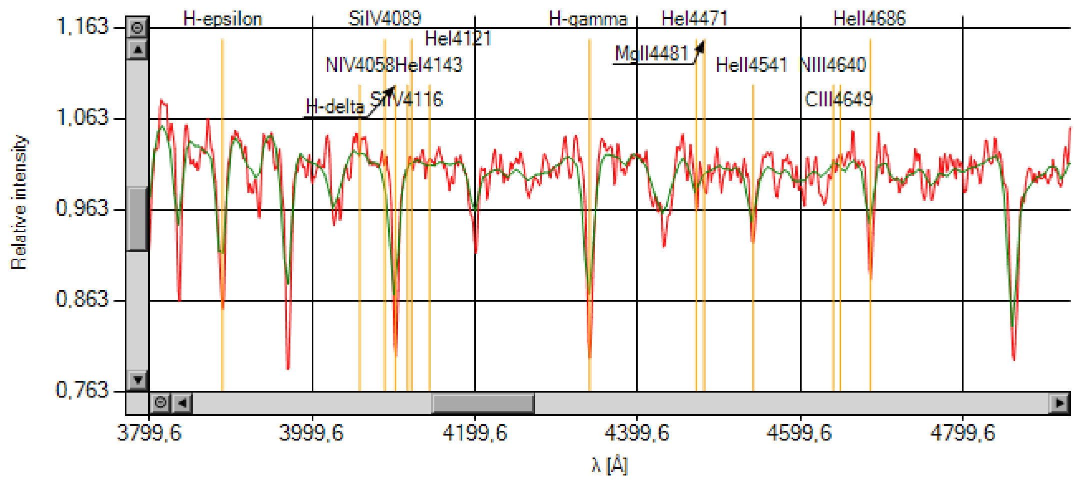

Figure II: Spectra of O5V stars: HD 242908 (red line) and MK O5V (green line). Modified method used to process spectra containing emission lines.

spectral line	W_{STAR} [Å]	$W_{STANDARD}$ [Å]
N IV λ4058	0.115	0.363
Si IV λ4089	0.065	not resolved
Si IV λ4116	0.077	not resolved
He I λ4121	0.043	0.238
He I λ4143	0.077	0.176
He I λ4471	0.125	not resolved
He II λ4541	0.293	0.528
Mg II λ4481	0.054	0.134
N III λ4640	0.055	0.180
Hϵ	1.120	1.008
Hδ	1.352	1.651
Hγ	1.547	1.663
C III λ4649	0.093	0.133
He II λ4686	0.622	0.504

(a) Table of important spectral lines and their equivalent widths measured in the spectra of HD 242908 and MK O5V.

criterion	W'_{STAR} [Å]	$W'_{STANDARD}$ [Å]	$\frac{W'_{STAR}}{W'_{STANDARD}}$
N IV λ4058 / N III λ4640	2.082	2.013	1.035
He I λ4471 / He II λ4541	0.425	0.254	1.675
Mg II λ4481 / He I λ4471	0.435	–	–
C III λ4649 / He II λ4686	0.149	0.263	0.568
Si IV λ4089 / He I λ4143	0.850	–	–
Si IV λ4089 / Hγ	0.042	–	–
Si IV λ4116 / He I λ4121	1.791	–	–
Hδ / He II λ4541	5.280	3.152	1.675

(b) Table of spectral line ratio criteria, their associated values and the ratio between the star's and the standard's criterion value.

Table IV: The criteria for classification of HD 242908. Modified method used to process spectra containing emission lines.

Star: HD 37767
Standard star: MK B3V

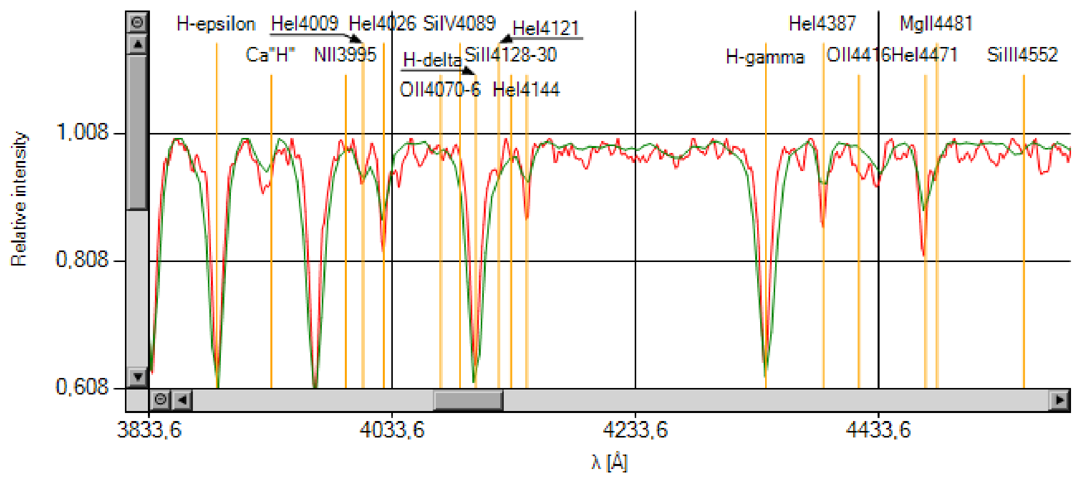

Figure III: Spectra of B3V stars: HD 37767 (red line) and MK B3V (green line).

spectral line	W_{STAR} [Å]	$W_{STANDARD}$ [Å]
Si III λ4552	0.055	0.141
Si IV λ4089	0.053	not resolved
He I λ4471	1.318	1.492
Mg II λ4481	0.252	not resolved
Si II λ4128-30	0.124	not resolved
Si IV + He I λ4121	0.233	0.577
He I λ4144	0.687	0.766
He I λ4026	1.174	1.510
O II λ4070-6	0.120	0.328
He I λ4387	0.802	0.875
O II λ4416	0.255	0.528
Hϵ	4.207	6.350
Hδ	3.961	7.070
Hγ	3.935	6.744
Hβ	4.083	7.062
Hα	0.204	5.524
N II λ3995	0.066	0.408
He I λ4009	0.268	1.517
Ca II "H"	0.338	0.637

(a) Table of important spectral lines and their equivalent widths measured in the spectra of HD 37767 and MK B3V.

criterion	W'_{STAR} [Å]	$W'_{STANDARD}$ [Å]	$\frac{W'_{STAR}}{W'_{STANDARD}}$
Si III λ4552 / Si IV λ4089	1.041	–	–
Mg II λ4481 / He I λ4471	0.191	–	–
Si II λ4128-30 / Si IV + He I λ4121	0.531	–	–
Si II λ4128-30 / He I λ4144	0.180	–	–
Si IV + He I λ4121 / He I λ4144	0.339	0.753	0.450
Si IV λ4089 / He I λ4009	0.199	–	–
He I λ4026 / O II λ4070-6	9.774	4.609	2.121
He I λ4387 / O II λ4416	3.147	1.657	1.900
N II λ3995 / He I λ4009	0.247	0.269	0.920
Hγ / Hβ	0.964	0.955	1.009
Hβ / Hα	20.042	1.278	15.677
Hδ / He I λ4471	3.099	4.735	0.654

(b) Table of spectral line ratio criteria, their associated values and the ratio between the star's and the standard's criterion value.

Table V: The criteria for classification of HD 37767.

Star: O 1015
Standard star: MK B8V

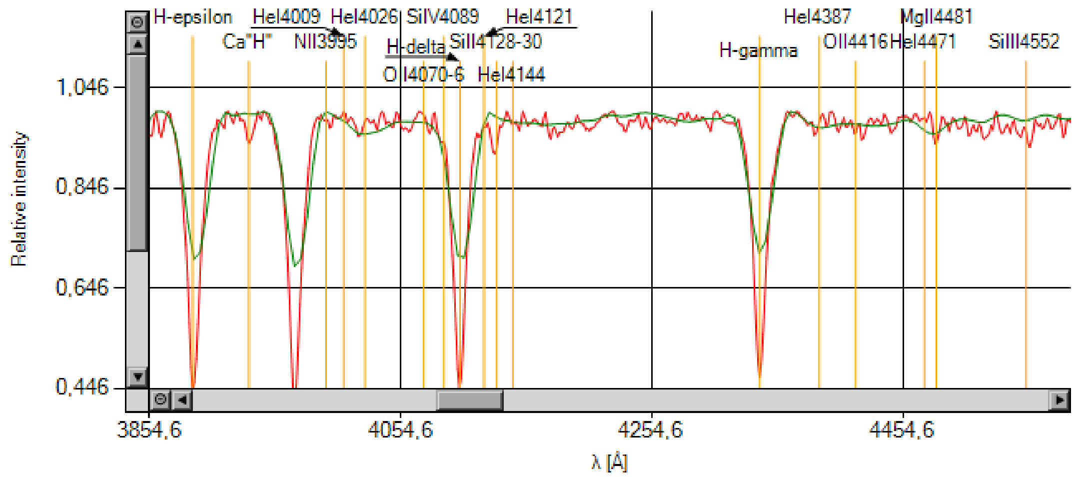

Figure IV: Spectra of B8V stars: O 1015 (red line) and MK B8V (green line).

spectral line	W_{STAR} [Å]	$W_{STANDARD}$ [Å]
Si III λ4552	0.157	0.258
Si IV λ4089	0.054	not resolved
He I λ4471	0.060	0.625
Mg II λ4481	0.239	not resolved
Si II λ4128-30	0.416	not resolved
Si IV + He I λ4121	0.126	3.407
He I λ4144	0.096	0.342
He I λ4026	0.072	0.665
O II λ4070-6	0.010	0.316
He I λ4387	0.041	0.471
O II λ4416	0.165	0.384
Hϵ	7.692	6.438
Hδ	7.022	6.608
Hγ	7.345	6.227
Hβ	7.547	5.577
Hα	5.263	4.675
N II λ3995	0.095	not resolved
He I λ4009	0.020	4.740
Ca II "H"	0.247	0.084

(a) Table of important spectral lines and their equivalent widths measured in the spectra of O 1015 and MK B8V.

criterion	W'_{STAR} [Å]	$W'_{STANDARD}$ [Å]	$\frac{W'_{STAR}}{W'_{STANDARD}}$
Si III λ4552 / Si IV λ4089	2.896	–	–
Mg II λ4481 / He I λ4471	3.963	–	–
Si II λ4128-30 / Si IV + He I λ4121	3.307	–	–
Si II λ4128-30 / He I λ4144	4.333	–	–
Si IV + He I λ4121 / He I λ4144	1.310	–	–
Si IV λ4089 / He I λ4009	2.753	–	–
He I λ4026 / O II λ4070-6	7.171	2.103	3.409
He I λ4387 / O II λ4416	0.246	1.227	0.201
N II λ3995 / He I λ4009	4.850	–	–
Hγ / Hβ	0.973	1.117	0.871
Hβ / Hα	1.434	1.193	1.202
Hδ / He I λ4471	125.230	8.925	14.031

(b) Table of spectral line ratio criteria, their associated values and the ratio between the star's and the standard's criterion value.

Table VI: The criteria for classification of O 1015.

Star: HD 221741
Standard star: MK A3V

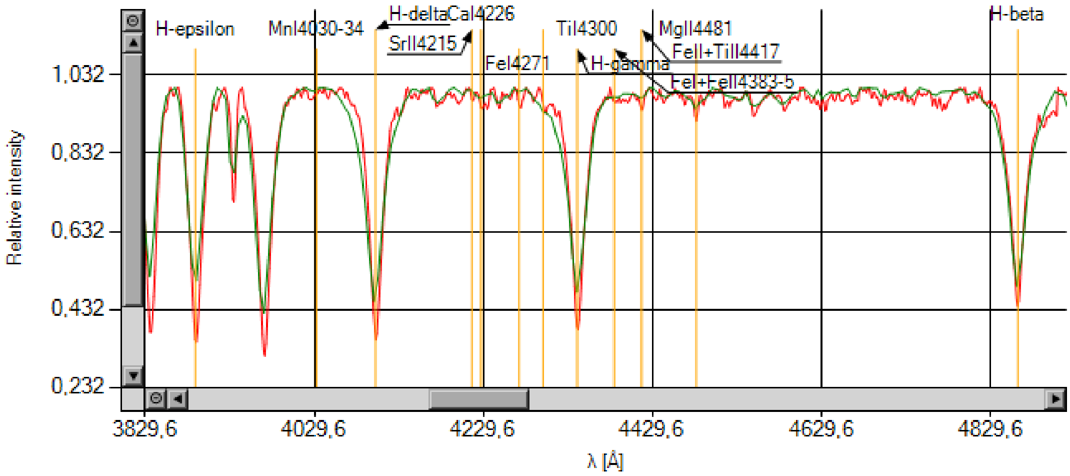

Figure V: Spectra of A3V stars: HD 221741 (red line) and MK A3V (green line).

spectral line	W_{STAR} [Å]	$W_{STANDARD}$ [Å]
Mn I λ4030-4	0.035	0.084
Ca I λ4226	0.212	0.329
Fe I λ4271	0.181	0.158
Ti I λ4300	0.243	not resolved
Mg I λI4481	0.315	0.612
Sr II λ4215	0.130	0.145
Hε	12.195	10.873
Hδ	13.262	16.270
Hγ	13.043	14.958
Hβ	13.065	15.317
Hα	7.858	9.584
Fe II + Ti II λ4417	0.214	0.331
Fe I + Fe II λ4383-5	0.146	0.269

(a) Table of important spectral lines and their equivalent widths measured in the spectra of HD 221741 and MK A3V.

criterion	W'_{STAR} [Å]	$W'_{STANDARD}$ [Å]	$\frac{W'_{STAR}}{W'_{STANDARD}}$
Ca I λ4226 / Hβ	0.016	0.021	0.754
Ca I λ4226 / Hγ	0.016	0.022	0.737
Mn I λ4030-4 / Hβ	0.003	0.006	0.488
Mn I λ4030-4 / Hγ	0.003	0.006	0.478
Fe I λ4271 / Hβ	0.014	0.010	1.348
Fe I λ4271 / Hγ	0.014	0.011	1.318
Ti I λ4300 / Hβ	0.019	–	–
Ti I λ4300 / Hγ	0.019	–	–
Fe II + Ti II λ4417 / Mg II λ4481	0.679	0.540	1.258
Mn I λ4030-4 / Hγ	0.003	0.006	0.478
Fe I + Fe II λ4383-5 / Mg II λ4481	0.462	0.440	1.051
Sr II λ4215 / Ca I λ4226	0.616	0.440	1.401
Hδ / Hγ	1.017	1.088	0.935
Hγ / Hβ	0.998	0.977	1.022
Hβ / Hα	1.663	1.598	1.040

(b) Table of spectral line ratio criteria, their associated values and the ratio between the star's and the standard's criterion value.

Table VII: The criteria for classification of HD 221741.

Discussion

O5V: HD 242908, MK O5V

The first tested spectra belongs to the O type main sequence star HD 242908, which was compared to the standard data of MK O5V. The spectra were processed using two different continua estimation methods. The first one, the standard peak stripping method, was aimed at determining the spectrum's absorption line profiles and equivalent widths, the second method was the modified peak stripping method with additional fit parameters, used to identify any present emission lines. Both cases show a solid corelation between the two spectra. The profiles of the spectral line of the tested star are quite sharper, than those of the standard, that is the manifestation of the difference between the spectrogram resolutions. This difference can be particularly observed in the evaluation of the very important Si IV lines, which were resolved in the spectrum of the tested star as opposed to the spectrum of the Pickles standard star. The Mg II spectral line remained unresolved in the spectra of MK O5V in both cases as well. The He I / He II ratio has been evaluated in all cases, as was the Hδ / H II. Both of these ratios increase with decreasing temperature and are therefore usable for the spectral part of stellar classification. For early O type stars, the ratio of N IV / N III, that was resolved in all cases too, can be additionally used. As for the luminosity classification, the C III / He II ratio carries the crucial information for O type stars. As the mentioned C and N spectral lines are mostly emission lines, the emphasis should be put on the results of the spectrum normalized by modified peak stripping method. If we assume that case, the relative difference between the tested star's and standard's ratios of N IV / N III amounts to 1.034, which supports the use of this criterion for spectral classification. As for the star's and standard's relative difference of C III / He II, the value of 0.568 is affected by the difference in spectrogram's resolutions as well as a possible difference in luminosities of the tested and the standard, or even their chemical composition, but as such still remains as a very important luminosity classification criterion. Further, the HD 242908 exhibits weak emission in N III, but no emission in He II and therefore does not qualify as a Oe star, neither does it exhibit emissions in balmer lines. It could be noted that main sequence stars most often do not exhibit strong emissions and such result may have been, to an extent, expected.

B3V: HD 37767, MK B3V and B8V: O 1015, MK B8V

The B type stars selected for measurement with SpecOp were HD 37767 and O 1015, while respective Pickles MK spectra were used as standards. Both tested stars were tested using the original peak stripping function and show a decent match, despite having slightly sharper profiles, in the case of B8V more so, possibly due to the forementioned spectrogram resolutions. Spectra of both tested stars had to be corrected for radial velocity shifts before processing, HD 37767 was corrected for z = 0.0022 and O 1015 for z = 0.0019. Spectral lines of both of the stars and both of the standards are dominated by the hydrogen absorption spectral lines, of which Hβ, Hγ, Hδ and Hε have profiles of similar strengths. Unfortunately, most of the Si spectral lines turned out to be too faint to be resolved in the spectra of the Pickles standard. Spectral classification criteria may be sought in the ratio of Hδ / He I, but the ratio is only reliable for the early B type stars. At the late B type stars, none of the spectral classification spectral lines is resolved in the Pickles spectrum, which leaves us empty-handed and we may have to consider using Hδ / He I, although this might not be the most reliable approach as He I tends to gradually disappear from the spectra at about this temperature. Enough criteria can be found in both tested stars and standards for the luminosity classification of B type stars, well resolved ratios of O II / He I, N II / He I for the early and Balmer line profiles for the early B type stars. Noteworthy is also the appearance of a blend at the position of Ca II "H" interstellar line, which appears in spectra of both HD 37767 and O 1015, and weakens toward the later types, until it changes into the true Ca II "H" spectral line later at the threshold between B and A type spectra.

A3V: HD 221741, MK A3V

The spectra of HD 221741 show solid match with the standard Pickles MK A3V. The Balmer lines dominate the spectra of both stars in all five measured balmer lines and in both spectra, the Hδ is the strongest measured spectral line. The ratios of the selected criterion ratios for each star show a correlation as well, the values range from 0.47 to 1.81, which is a rather small interval around the value of 1, if considering how different can stars even of the same subtype and luminosity class be. The only line remaining unresolved, Ti I 4300, belongs to the standard MK A3V. In the spectra, we can observe a very strong Ca II "H" absorption line that stays strong until later spectral types. Both Fe I and Ca I lines were resolved and possibly can be used for spectral classification of most subtypes of A type stars. For the luminosity classification, spectral lines Ca I, Sr II, Mg II and the blend of Fe II and Ti II were resolved in both spectra and are therefore valid criteria for all A subtypes.

Conclusions

In this thesis I have compiled a set of stellar classification criteria applicable to O, B and A type main sequence stars and developed standalone software capable of processing stellar spectra and evaluating equivalent widths of important spectral lines. Further, I have shown what stellar classification criteria can be used in conjunction with the software and universal Pickles MK standards. It was also shown that particular difficulty may arise when seeking equivalent widths of the important Si IV and Mg II spectral lines in the low resolution standard spectra, but usable criteria for classification can be found for nearly any spectral subtype, except for the late B type stars, even after the exclusion of lines that were not resolved in the standard spectra.

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