Astronomía Avanzada I (Semester 1 2025)

Stellar Atmospheres (1)

Introduction to Stellar Atmospheres

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May 23, 2025

Introduction

course logistics:

- first two lectures: stellar atmospheres, last two lectures: variable stars
- lecture dates: May 23, May 27, May 30, June 3, June 6
- paper presentation (of your choice) on June 6

contact and course material:

- e-mail: `nina.hernitschek@uantof.cl`
- github: `https://github.com/ninahernitschek/astronomia_advanzada_I_2025_1`

Introduction

paper presentation:

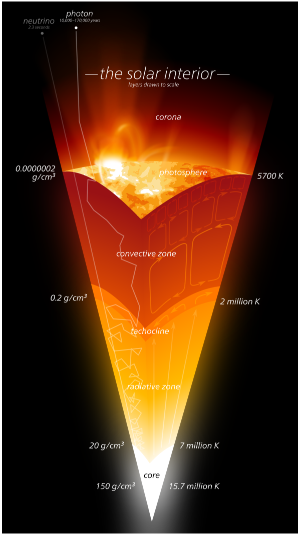
- 20 minutes
- LaTeX
- figures: properly cite and describe figures from the paper
- data of project shown in the paper: data description (incl. citation)
- discussion and summary: reflect the approach (strengths, weaknesses, limitations), lessions learned
- bibliography: bibtex/ref mechanism, ADS/Bibtex information

Introduction

In the following lectures, you will learn about the physics of stellar spectra including spectral lines, and radiation transport occuring in stellar atmospheres.

What are Stellar Atmospheres?

The stellar atmosphere is what you can actually see of a star!



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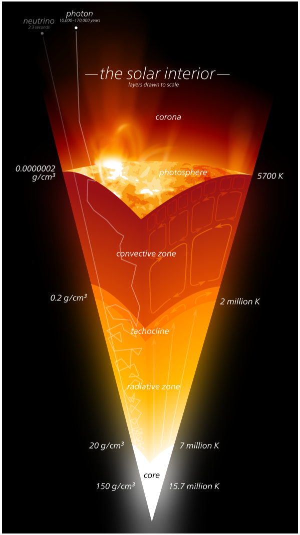
We divide stars into two major sectors: the **stellar interior** (inaccessible to photometric and spectroscopic observation*) and the **atmosphere** (the visible portion of the star and the transition between interior and exterior).



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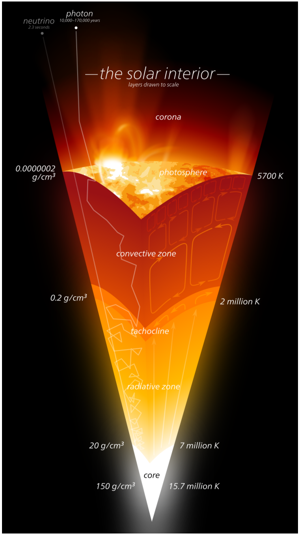
In contrast to the interior, where **convection** may dominate, the energy transport mechanism in the atmospere is **radiation**.

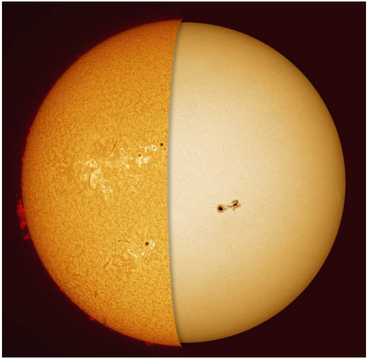
* The stellar interior can be probed with Astereoseismology, the study of the internal structures of stars by means of their intrinsic global oscillations.



Features of Stellar Atmospheres

The photosphere is the visible disk, while the atmosphere also includes coronae and winds.

Hα light is emitted by hydrogen atoms. Their electrons absorb energy and rise to a higher energy level. When cascading back to their original level, they release that energy as light with a wavelength of 6562.8 Å (656 nm).

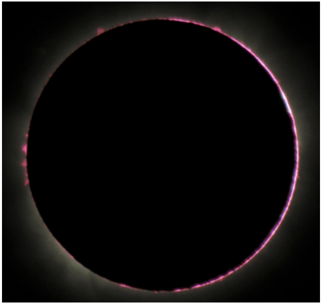


Composite image of the Sun in Hα (left) and white light (right). In Hα, we see the solar chromosphere, the layer directly above the photosphere. (For that reason, in Hα the Sun appears larger.) credit: Alan Friedman/ avertedimagination.com

Features of Stellar Atmospheres

The 1,250-mile-thick layer of solar atmosphere located just above the photosphere is called the **chromosphere**.
Here, the temperature rises from about 6370 K to 20260 K, hot enough to excite H to emit its singular red light.

Chromo means color and refers to the red **prominences** seen around the limb of the Sun when its overly-brilliant photosphere is covered up by the Moon in a total solar eclipse or by using a Coronagraph and an Hα filter.

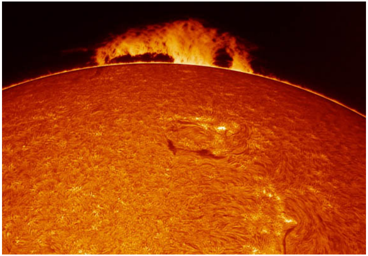


The Sun's crimson-hued chromosphere and prominences were captured during the November 2013 total solar eclipse from Pokwero Village, Uganda.
credit: Alson Wong

Features of Stellar Atmospheres

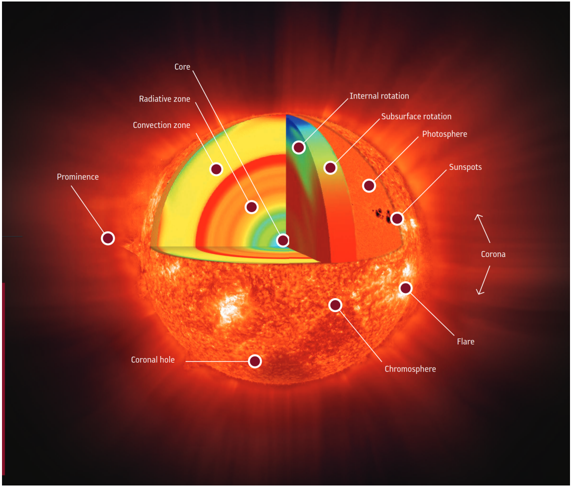
A **plage** is a bright region in the Sun's chromosphere, typically found in and around active regions. Historically, they have been referred to as *bright flocculi*.

Classically as regions that are bright in Hα and other chromospheric emission lines, nowadays most researchers identify plage based on the photospheric magnetic field concentration of the faculae below.
It is believed that plage is formed from decaying emerging flux regions, and often acts as a footprint for coronal loops.



A prominence photographed through a Hα scope on September 17, 2015. Bright white patches are "plages". credit: Bob Antol / stargate4173.com

Features of Stellar Atmospheres



Cutaway of the Sun showing its core, where nuclear fusion occurs, radiative zone, convection zone and the three layers of its atmosphere: photosphere, chromosphere and corona. credit: NASA

Why we care about Stellar Atmospheres?

About 2/3 of the light is **absorbed** in the atmosphere.

Stellar interiors are effectively invisible to external observers (apart for e.g. astroseismology) so all the information we receive from stars originates from their atmospheres.

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Stellar atmospheres are primarily characterized by two parameters: T_{eff}, $\log g$.

The transition between the stellar inerior and the interstellar medium is seen in the change of the **physical parameter profile** (temperature, surface gravity $g = g_{\odot} M/R^2$ or density).

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The **behavior** of stellar atmospheres depends mainly on the mass and age of the star and, secondarily, on the chemical composition, angular momentum and magnetic field.

Understanding Stellar Atmospheres

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Important for these are:

- plasma physics (e.g. line broadening)
- atomic physics (microscopic interaction between light and matter)
- radiative transfer (macroscopic interaction between light and matter)
- thermodynamics
- stellar properties
- chemical composition

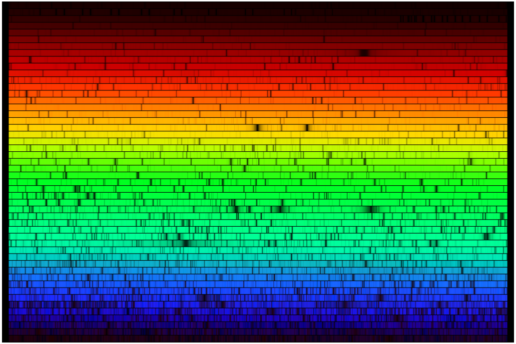
Observing Stellar Atmospheres

As we've seen: observing stars is (automatically) observing stellar atmospheres!

But we can tailor our observation towards getting more relevant data for understanding stellar atmospheres.

Spectra as Continuous Energy Distributions

Spectra show us how energy emission is distributed over a range of wavelengths.

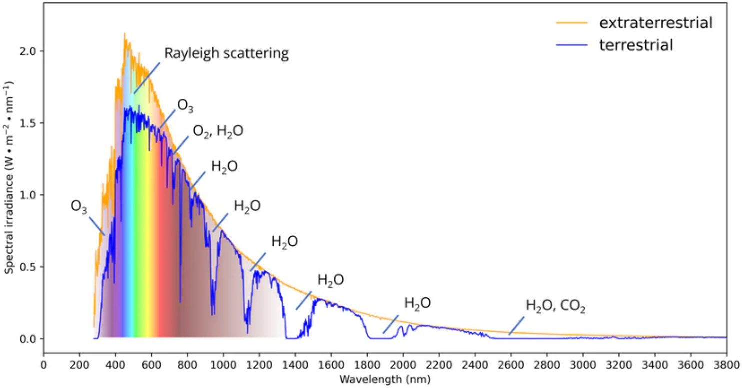


A 2D Echelle image of the optical Solar spectrum. Credit: N. A. Sharp.

However, in order to study a spectrum in detail (to see subtle differences in brightness of different colors) it needs to be plotted as a graph.

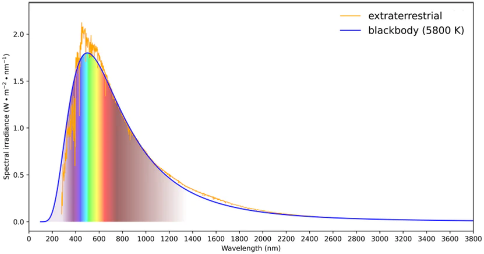
Spectra as Continuous Energy Distributions

Absorption and scattering change the spectral distribution of sunlight as it passes through the atmosphere.



Spectra as Continuous Energy Distributions

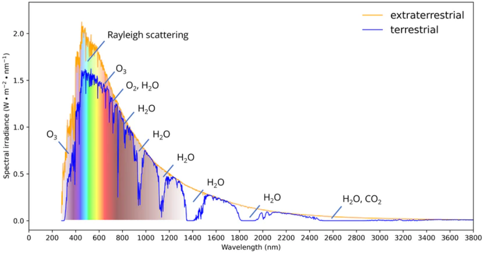
Observationally we differentiate two important components: **Continuum and spectral lines**.



The continuum can be described as a **black body**. Emission/absorption lines respond to interactions between particles (photons-atoms/molecules).

Spectra as Continuous Energy Distributions

Absorption is the process where a photon transfers all its energy to an atom/molecule.



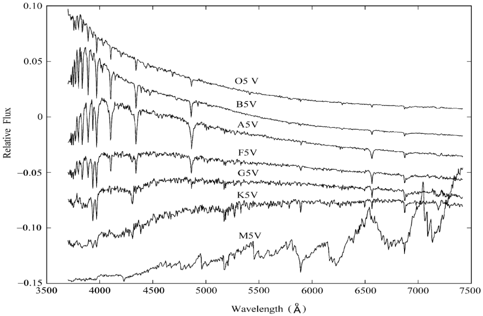
At UV wavelengths below 400 nm, O_3 (ozone) is the primary absorber.

Molecules in the atmosphere **scatter** incoming solar radiation into random directions. This is known as Rayleigh scattering and is responsible for the reduction in irradiance throughout the visible part of the spectrum.

Local effects, such as weather and pollution, can further change the terrestrial spectrum.

Spectra as Continuous Energy Distributions

Stellar spectra can differ a lot:



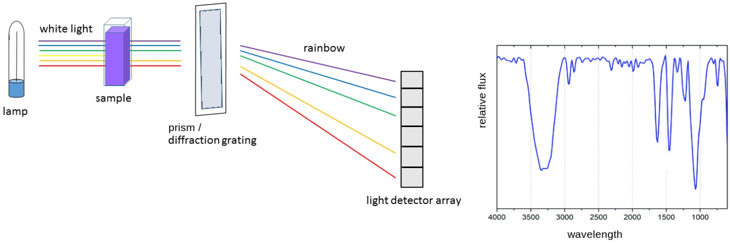
Seven main types of stellar spectra. Credit: D.-M. Qin et al. (2003)

We will later see that stellar spectra are an important tool to **classify stars**.

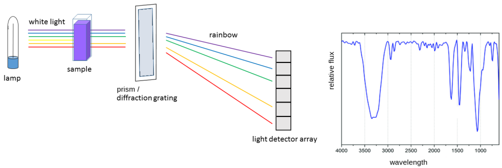
Spectroscopy

Isaac Newton demonstrated that white light can be separated into its component colors.

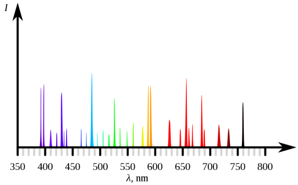
A **spectroscope** consists of a prism or grating to split up light into its colors (wavelengths) and then projects them onto a screen or detector for analysis, producing a continuous spectrum.

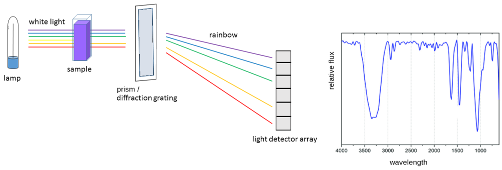


Spectroscopy

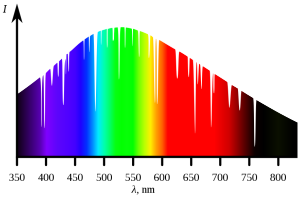


When examining light from thin, **hot gas**, we see a series of discrete bright lines, the **emission lines**. They are produced when electrons drop from an excited state to a lower state. The wavelengths emitted are unique for each element (Bohr model). Thus, by examining the emission lines, we can determine the elements in the gas.





When examining light passing through a **cool gas**, we see a continuum with blank lines in which specific wavelengths have disappeared - the **absorption lines**. They are the result of electrons absorbing light. The specific wavelengths in the absorption lines of an element are the same as the wavelengths in the emission lines of the same element.
Thus, we can identify elements by either their emission or absorption spectra.



Kirchhoff Radiation Laws

The radiation laws of Gustav Kirchhoff ($\sim$1840) describe the phenomenon:

1. A hot solid, liquid, or dense gas produces a continuous spectrum.

2. A hot diffuse gas under low pressure produces an emission line spectrum.

3. A cold diffuse gas under pressure in front of a continuous spectrum source produces dark spectral lines (absorption lines) in the continuous spectrum.

Although these laws were developed in **laboratories**, they describe the emission phenomena associated with the gas and are fundamental for describing atmospheres.

Astronomy with Spectra

Despite initially developed for the H atom, the mechanism works as well for multielectron atoms and even molecules, who have more complex spectra.

All these spectra behave in predictable fashion, so each acts a definitive "fingerprint" that can be used to identify the element or molecule.

Kirchhoff's Laws have proven to be useful in identifying **composition and temperature** of e.g. planetary atmospheres, stars, and interstellar nebula.

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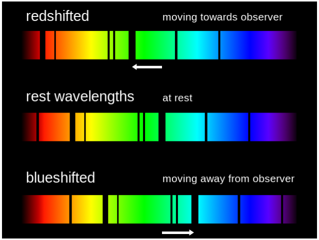
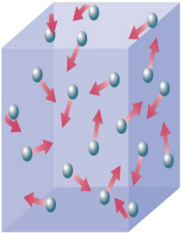
Indirectly, from spectra also the motion of those objects can be inferred, using the **Doppler effect**.



Black-Body Radiation

All baryonic matter (above a temperature of absolute zero) emits electromagnetic radiation.

The radiation represents a conversion of a body's internal energy into electromagnetic energy, and is therefore called **thermal radiation**.



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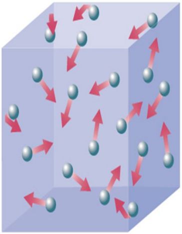


Black-body radiation is the thermal electromagnetic radiation within, or surrounding, a body in thermodynamic equilibrium with its environment, emitted by an idealized black body (opaque, non-reflective).

Black-body radiation was first observed by Th. Wedgwood (1792).

Black-Body Radiation

Subsequent investigations conclude that:

1. Every body with temperature emits light at all wavelengths with varying efficiency.

2. An ideal emitter absorbs all the light incident on it and re-radiates it.

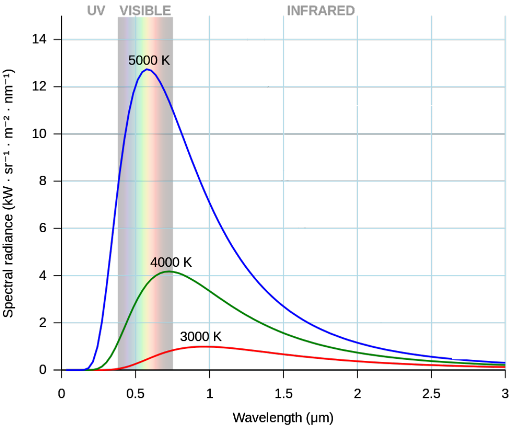
3. The wavelength of maximum radiation is given by **Wien's displacement law**:

The spectral radiance of black-body radiation per unit wavelength, peaks at the wavelength $\lambda_{\max}$ given by:

$$\lambda_{\max} = \frac{b}{T}$$

where T is the absolute temperature and b is a constant of proportionality called Wien's displacement constant, $b = 2.897771955 \times 10^{3}\ \mathrm{m\,K}$.

Black-Body Radiation



Black-body radiation as a function of wavelength for various temperatures. Each temperature curve peaks at a different wavelength, as described by Wien's law. With increasing temperature, the peak of the blackbody radiation curve moves to higher intensities and shorter wavelengths.

Black Body Radiation

Not only the wavelength at which the spectral radiance peaks shifts with temperature:

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For a black body, the **Stefan-Boltzmann law** states that the total energy radiated per unit area per unit time (also known as the flux) is directly proportional to the fourth power of the black body's temperature:

$$F = \sigma_{\mathrm{SB}}\, T^4.$$

The constant of proportionality, σ_{SB}, is called the Stefan-Boltzmann constant. It has a value $\sigma_{\mathrm{SB}} = 5.670374419 \times 10^{-8}\,\mathrm{W\,m^{-2}\,K^{-4}}$.

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We can use the Stefan-Boltzmann law to calculate the **star's temperature**: When approximating a star as black body, the total luminosity of a star is $L = 4\pi R^2 \sigma_{\mathrm{SB}} T^4$.

example: For the Sun, $T \sim 6000\ \mathrm{K}$, $R_\odot = 6.96 \times 10^8\ \mathrm{m}$, so $L_\odot \sim 4 \times 10^{26}\ \mathrm{W}$.

Electron Transitions

We've seen that:

- emission lines are produced when electrons drop from an excited state to a lower state
- absorption lines are the result of electrons absorbing light, jumping up to a higher state

Electron Transitions

Observations of hydrogen spectra lead to the following distribution of hydrogen lines (Johann Balmer (1885) and Johannes Rydberg (1888)):

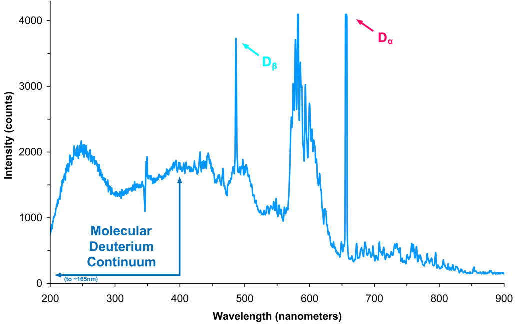
$$\frac{1}{\lambda} = R_H \left(\frac{1}{2^2} - \frac{1}{n^2} \right) \quad \text{for } n = 3, 4, 5, \ldots$$

where λ is the wavelength of the absorbed/emitted light and R_H is the Rydberg constant for hydrogen, $R_H = 1.09677 \times^7 \ \mathrm{m}^{-1}$.

The "visible" hydrogen emission spectrum lines in the Balmer series. Hα is the red line at the right. Four lines (counting from the right) are formally in the visible range. Lines five and six can be seen with the naked eye, but are considered to be UV as they have $\lambda < 400$ nm.

Electron Transitions

Two of the Balmer lines (α and β) are clearly visible in this emission spectrum of a deuterium lamp.

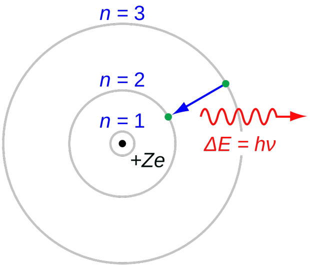


Deuterium (hydrogen-2, symbol ^2H or D, also known as *heavy hydrogen*) is one of two stable isotopes of hydrogen (the other is protium, or hydrogen-1). In addition to one proton, it contains one neutron in its nucleus.

Electron Transitions

In the simplified Rutherford-Bohr model of the H atom, the Balmer lines result from an electron jump between the second energy level closest to the nucleus, and those levels more distant.

Shown here is a photon emission. The $3 \rightarrow 2$ transition depicted here produces $H\alpha$, the first line of the Balmer series. For H, this transition results in a photon of wavelength 656 nm (red).



Electron Transitions

After Balmer's discovery, five other hydrogen spectral series were discovered, corresponding to electrons transitioning to values of n other than two.

We find a **generalized** equation for the wavelengths of emitted or absorbed photons:

$$\frac{1}{\lambda} = Z^2 R_\infty \left(\frac{1}{{n_1}^2} - \frac{1}{n_2^2} \right)$$

where

Z is the atomic number
n_1 is the principal quantum number of the lower energy level
n_2 is the principal quantum number of the upper energy level
R_∞ is the Rydberg constant ($1.09677 \times 10^7\ \mathrm{m}^{-1}$ for hydrogen and $1.09737 \times 10^7\ \mathrm{m}^{-1}$ for heavy metals).

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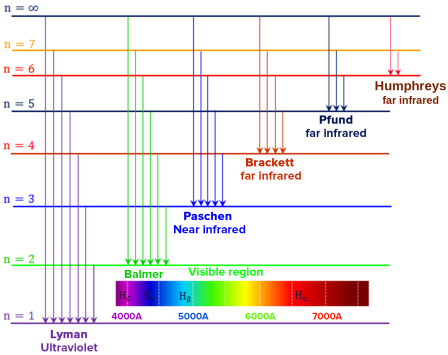
This equation is valid for all atoms having only a single electron, and the particular case of hydrogen spectral lines is given by Z=1.

Electron Transitions

The transitions are named the following:

$n_1 = 1$: Lyman
$n_1 = 2$: Balmer
$n_1 = 3$: Paschen
$n_1 = 4$: Brackett
$n_1 = 5$: Pfund
$n_1 = 6$: Humphreys



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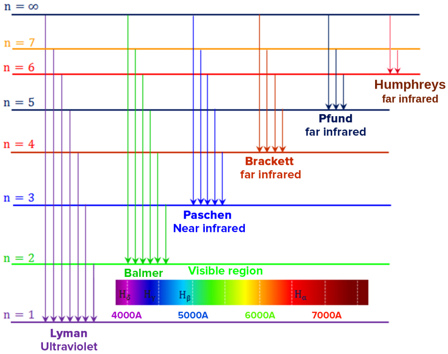

Further series ($n_1 > 6$) are unnamed, but follow the same pattern and equation as dictated by the Rydberg equation. Series are increasingly spread out and occur at increasing wavelengths. The lines are also increasingly faint, corresponding to increasingly rare atomic events.

Balmer Lines in Astronomy

Because the Balmer lines are commonly seen in the spectra of various objects, they are often used to **determine radial velocities due to Doppler shifting** of the Balmer lines.

Common **use cases** are such as:

- measuring stellar pulsations
- detecting binary stars
- detecting exoplanets
- detecting compact objects such as neutron stars and black holes (by the motion of Hydrogen in accretion disks around them)
- identifying groups of objects with similar motions and presumably origins (moving groups, star clusters, galaxy clusters)
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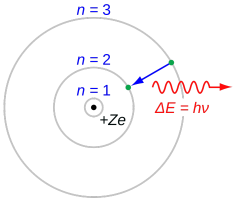
In stars, the Balmer lines are usually seen in **absorption**, and they are "strongest" in stars with a surface temperature of about 10,000 K (spectral type A). In the spectra of most spiral and irregular galaxies, AGN, H II regions and planetary nebulae, the Balmer lines are **emission lines**.

Atom Models and Spectral Lines

Spectral lines can be conveniently explained with the Bohr-Rutherford model in which spectral lines result from electron jumps between levels.



However, this model fails to explain certain effects, such as:

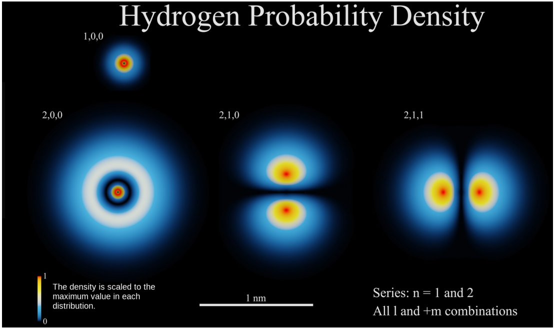
- The relative intensities of spectral lines; although in some simple cases, Bohr's formula or modifications of it, was able to provide reasonable estimates.
- Doublets and triplets appear in the spectra of some atoms as very close pairs of lines. Bohr's model cannot say why some energy levels should be very close together.

Atom Models and Spectral Lines

In modern **quantum mechanics**, the electron in Hydrogen is a spherical cloud of probability that grows denser near the nucleus.

The electron and nucleus are viewed as probability functions, defining orbits in which the electron is most likely to be found:

The state of the particle becomes defined by a series of discrete **quantum numbers** (n, l, m_l and m_s), quantizing energy, angular mmentum and spin.



Parameters for Describing Stellar Atmospheres

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The **effective temperature** (in K) is defined by $L = 4\pi R^2 \sigma T_{\rm eff}^4$, which is related to **ionization**.

The **surface gravity** (in cm/s^2), $g = GM/R^2$, is related to **pressure**.

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example:
The Sun has $T_{\rm eff} = 5777\ \rm K$, $\log g = 4.44$. Its atmosphere is only a few hundred km deep, $<$0.1% of the stellar radius.
The Solar atmosphere is most easily studied during total eclipse.
A red giant has $\log g \sim 1$ (extended atmosphere), while a white dwarf has $\log g \sim 8$ (effectively zero atmosphere), and neutron stars have $\log g \sim 14 - 15$.

Parameters for Describing Stellar Atmospheres

?

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Parameters for Describing Stellar Atmospheres

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The atmospheric parameters are directly related to the stellar mass (M), radius (R) and luminosity (L).

A **fundamental star** has at least one of its atmospheric parameters obtained without reference to model atmospheres. An ideal fundamental star will have both parameters measured. These stars are vital for the quality assurance of model predictions.

Parameters for Describing Stellar Atmospheres

? How do we measure $T_{\rm eff}$ and $\log g$?

Unfortunately, the direct determination of $T_{\rm eff}$ and $\log g$ is not possible for most stars. Hence, we have to use **indirect methods**.

In most cases, **model atmospheres** are our analytical link between the physical properties of the star and the observed flux distribution and spectral line profiles. These observations can be used to obtain values for the atmospheric parameters $T_{\rm eff}$ and $\log g$, assuming of course that the models used are adequate and appropriate.

Parameters for Describing Stellar Atmospheres

The **effective temperature** of a star is physically related to the total radiant power per unit area at stellar surface (F_*):

$$\sigma T_{\mathrm{eff}}^4 \equiv \int_0^\infty F_\nu d\nu = F_* = \frac{L}{4\pi R^2}$$

It is the temperature of an equivalent **black body** that gives the same total power per unit area, and is directly given by stellar luminosity and radius. Since there is no true **surface** of a star, the stellar radius can vary with the wavelengths of observation and the nature of the star.

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The **stellar angular diameter** θ can be obtained directly using techniques such as speckle photometry, interferometry, and lunar occultations, and indirectly from eclipsing binary systems with known distances.

Parameters for Describing Stellar Atmospheres

The **surface gravity** of a star is directly given by the stellar mass and radius (in solar units):

$$g = g_{\odot} \frac{M}{R^2}$$

or, logarithmically,

$$\log g = \log M - 2 \log R + 4.437$$

Surface gravity is a measure of the photospheric pressure of the stellar atmosphere. Direct measurements are possible from eclipsing spectroscopic binaries, but again be aware of hidden model atmosphere dependences.

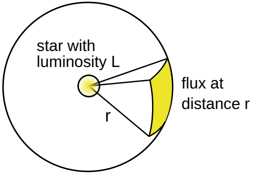
Flux

Flux (or radiant flux), F, is the total amount of energy that crosses a unit area per unit time. Flux is usually given in watts per square meter ($\mathrm{W/m^2}$).

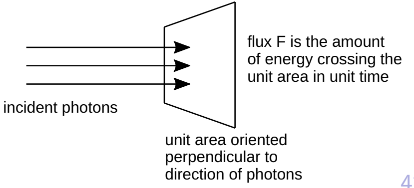
The flux of an astronomical source depends on the luminosity of the object and its distance from the Earth, according to the inverse square law:

$$F = \frac{L}{4\pi r^2}$$

where F = flux measured at distance r,
L = luminosity of the source,
r= distance to the source.



surface area of sphere $4\pi d^2$



Flux

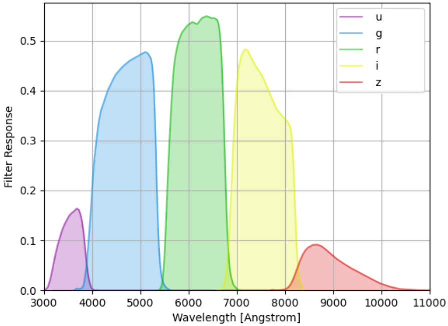
example: The luminosity of the Sun is $L_\odot = 3.839 \times 10^{26}\ \mathrm{W}$.
At a distance of $1\ \mathrm{AU} = 1.496 \times 10^{11}\ \mathrm{m}$, Earth receives a radiant flux above its absorbing atmosphere of

$$F = \frac{L}{4\pi r^2} = 1365\ \mathrm{W\,m^{-2}}.$$

Magnitudes and Broad-Band Filters

We often measure the flux F from astronomical objects via a logarithmic magnitude scale.

Magnitudes almost universally involve a set of **broad-band filters**, e.g. Johnson *UBVRI* or Sloan *ugriz*:



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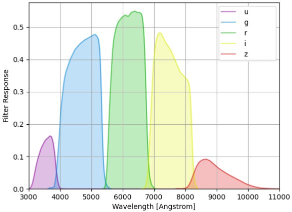

We then calculate:

$$m = -2.5 \log \int_0^\infty F_\nu W(\nu)\,\mathrm{d}\nu + \mathrm{const}$$

with:
F_ν a star's spectral energy distribution (SED)
$W(\nu)$ a filter passband

Color Index

We can define a **color index** as the difference between filters relative to Vega, e.g.

$$B - V = m_B - m_V = -2.5 \log \left(\frac{\int F_\nu W_B(\nu) \mathrm{d}\nu}{\int F_\nu W_V(\nu) \mathrm{d}\nu} \right) + 0.710$$

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example: $(B - V)_\odot = +0.65$ mag.

Absolute Magnitude

We can define the **absolute (visual) magnitude** M_V as the apparent (visual) magnitude of a star m_V at a distance of $d = 10\ \mathrm{pc}$.

Because $F \propto d^{-2}$:

$$\begin{aligned} M_V - m_V &= -2.5 \log\left[F(10\ \mathrm{pc})/F(d)\right] \\ &= -5 \log(d/10\ \mathrm{pc}) \\ &= 5 - 5 \log(d/\mathrm{pc}) \end{aligned}$$

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Visually, the **extinction term** is $A_V \sim 3.1E(B - V)$ for most sight lines, where $E(B - V) = -V - (B - V)_0$, i.e. the difference between observed and intrinsic $B - V$ color.
Interstellar extinction is much higher at shorter wavelengths, so IR observations of e.g. the Milky Way disk probe much further.

Bolometric Flux

The **bolometric flux** (erg cm^{-2} s^{-1}) from a star received at the top of Earth's atmosphere is the integral of the spectral flux (measured at a frequency ν or wavelength λ) over all the frequencies or wavelengths:

$$F_{\mathrm{bol}} = \int_0^\infty F_\nu \,\mathrm{d}\nu = \int_0^\infty F_\lambda \,\mathrm{d}\lambda$$

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Since the Earth's atmosphere is opaque to UV and some IR radiation, one cannot always directly measure the bolometric flux.

Bolometric Flux

The **bolometric correction** is the correction made to the absolute magnitude of an astronomical object in order to convert its visible magnitude to its bolometric magnitude.
From (Earth) atmosperic models, one can calculate **bolometric corrections** (BC) to correct measured fluxes, usually in the V band, for the total (bolometric) flux.
Usually the bolometric correction is expressed in magnitudes:

$$BC = M_{\mathrm{bol}} - M_V$$

with $M_{\mathrm{bol}} = 4.74 - 2.5 \log(L/L_\odot)$
The bolometric correction scale is set by the absolute magnitude of the Sun and an adopted (arbitrary) absolute bolometric magnitude for the Sun.

example:
For our Sun, BC =-0.08 mag. This is a small correction since our Sun emits most radiation in the visual.
Hot OBs stars have a very large negative BC, since most of the energy is emitted in the UV.

Bolometric Flux

The bolometric correction for a range of stars with different **spectral types** and groups is given in the following table:

Spectral Type	Main Sequence	Giants	Supergiants
O3	-4.3	-4.2	-4.0
G0	-0.10	-0.13	-0.1
G5	-0.14	-0.34	-0.20
K0	-0.24	-0.42	-0.38
K5	-0.66	-1.19	-1.00
M0	-1.21	-1.28	-1.3

The bolometric correction is large and negative both for early type (hot) stars and for late type (cool) stars: For early type stars, a substantial part of the produced radiation is in the UV. For late type stars, a large part is in the IR. For a star like the Sun, the correction is only marginal because the Sun radiates most of its energy in the visual wavelength range.

Bolometric corrections are typically **derived** from using a color-magnitude diagram and a theoretical model of the star's atmosphere.

Spectral Classes

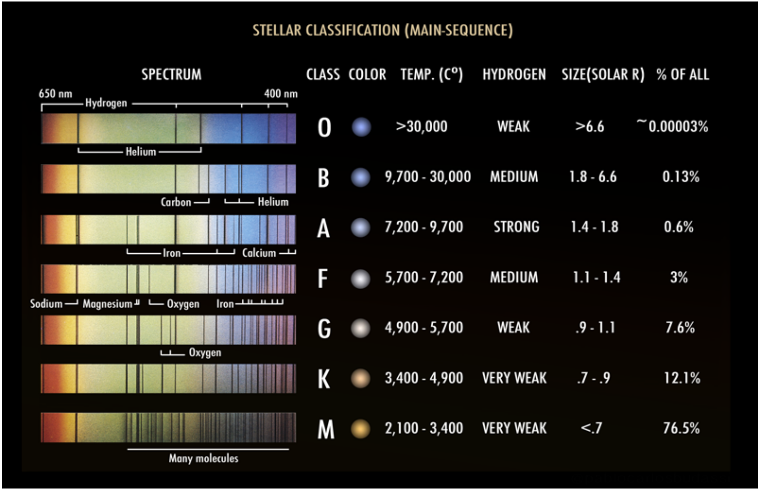
The **spectral classes** (Morgan-Keenan (MK); earlier version called the Harvard classification scheme) uses **ratios of spectral line strengths** to classify stars.

MK spectral class	class characteristics
O	hot stars with He II absorption
B	He I absorption; H developing later
A	very strong H, decreasing later; Ca II increasing
F	Ca II stronger; H weaker; metals developing
G	Ca II strong; Fe and other metals strong; H weaker
K	strong metallic lines; CH and CN bands developing
M	very red; TiO bands developing strong

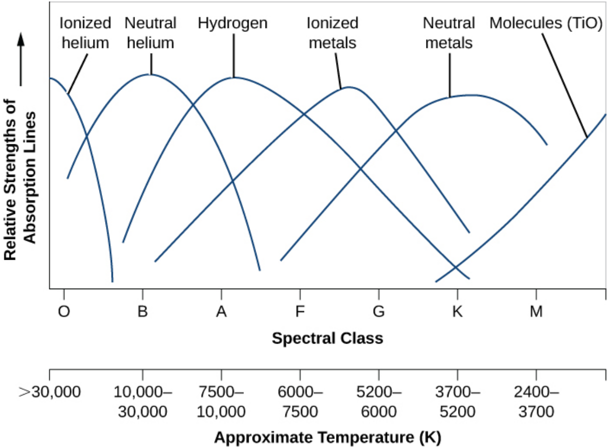
O-type stars have the bluest $B - V$ and highest T_{eff}. OBA stars are *early-type* stars, whereas cooler stars are *late-type*.

Spectral classes are each subdivided into up to 10 divisions, e.g. O2 ... O9, B0 ... B9

Spectral Classes



Spectral Classes



Luminosity Classes

There are certain problems with the spectral classes:

- spectra could have the same lines present (indicating same effective temperature), but different line widths
- stars could have the same effective temperature (hence also color and spectral class) but vary enormously in luminosity and thus absolute magnitude.

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To account for this, a second classification scheme, the **Luminosity Class** was added to the original concept of Spectral Class.

Luminosity Classes

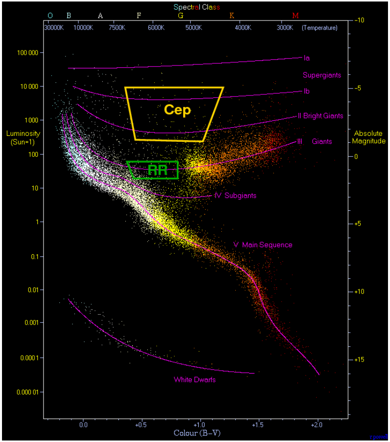
Luminosity class	class characteristics
0 or Ia+	hypergiants or extremely luminous supergiants
Ia	luminous supergiants
Iab	intermediate-size luminous supergiants
Ib	less luminous supergiants
II	bright giants
III	normal giants
IV	subgiants
V	main-sequence stars (dwarfs)
sd (prefix) or VI	subdwarfs
D (prefix) or VII	white dwarfs

Marginal cases are allowed; for example, a star may be either a supergiant or a bright giant, or may be in between the subgiant and main-sequence classifications. In these cases, two special symbols are used:

A slash (/) means that a star is either one class or the other.
A dash (-) means that the star is in between the two classes.

The Hertzsprung-Russel (H-R) Diagram

early 20th century: E. Hertzsprung and H. N. Russell noted that a scatter plot of a measure of stellar **luminosity** (absolute magnitude) vs. a measure of the stellar **temperature** (spectral type, color) shows structure:



An observational Hertzsprung-Russell diagram with 22,000 stars from the Hipparcos Catalogue and 1,000 from the Gliese Catalogue of nearby stars.

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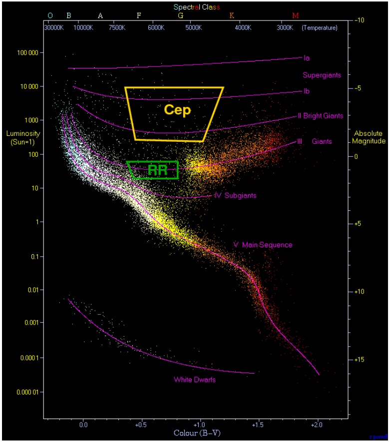

main sequence: a band extending from hot, high-luminosity stars to cool, low-luminosity stars, containing the vast majority of stars; Hydrogen fusion (H burning, H $\rightarrow$ He) in the core; the longest evolution stage

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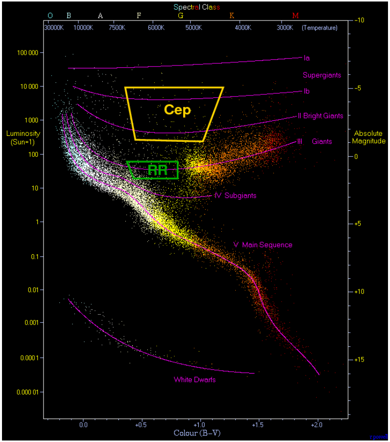

giants: After $\sim 10\%$ of the mass of a $M < \mathrm{M}_{\odot}$ star has been converted from H to He, the star expands: it becomes a red giant with a helium core and a hydrogen buring shell. The star heats up and moves along the horizontal branch in the H-R diagram, then cools off as the core burning stops. The once again red star then goes to higher luminosity on the asymptotic giant brach (AGB). Finally it ejects its envelope and becomes a planetary nebulae containing a white dwarf ($M \leq 1.4\mathrm{M}_{\odot}$, $R\sim\mathrm{R}_{\oplus}$).

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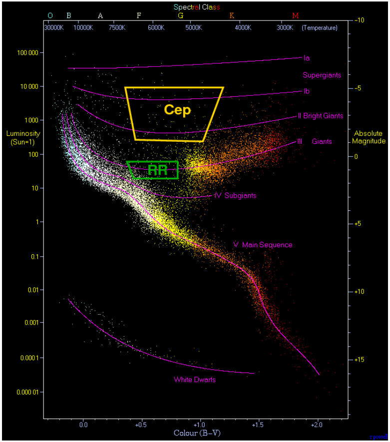

instability strip:
Stars, such as RR Lyrae stars and Cepheids, that have evolved off the main sequence and pulsate due to He III (doubly ionized helium). When the star contracts, the density and temperature of the He II layer below the atmosphere increases. He II starts to transform into He III (second ionization). This causes the opacity of the star to increase, the energy flux from the interior of the star is effectively absorbed. The temperature of the star rises and it begins to expand. After expansion, He III begins to recombine into He II, the opacity drops. This lowers the surface temperature of the star. The outer layers contract and the cycle starts from the beginning.

Summary

The stellar atmosphere is a **region** connecting unseen interior to vacuum exterior, and includes photosphere, chromosphere and corona.

The stellar atmosphere is **characterized** by ($T_{\rm eff}$, log g) which links physical properties of atmosphere to spectral classification (Spectral Type, Luminosity class).

Primary energy transport mechanism: radiation.

Stellar flux measured in magnitudes via broad-band filters, using Vega as zero point.

Bolometric flux represents integral over all wavelengths (related to visual flux via bolometric correction).