Properties Governing Lines and Continuum (mostly lines)

1. Metallicity / Abundances
   - number of absorbers/emitters
   - column density
   - electron density

2. Temperature I
   - thermal broadening (Doppler)
   - curve of growth (flat part)

3. Microturbulence
   - turbulence broadening

4. Temperature II
   - radiation field (Planck function)
   - excitation and ionization balance

5. Surface Gravity
   - pressure gradient (densities and line broadening)

6. Stellar Rotation
   - macroscopic broadening of lines

---

most important

- Abundances
- Temperature
- Gravity

---

Endary

- Microturbulence
Stellar Properties from Spectral Lines

Introduction

The spectral lines of a star contain a wealth of information, because their formation is sensitive to the conditions present in the star's atmosphere. For the astronomer this is simply fabulous, because it means that the light from a star can tell us a lot about what's going on!

Composition:
The spectral lines are formed by the absorption of photons as electrons jump from a lower energy level to a higher one. The specific wavelength (color) of light which is absorbed depends on the size of the jump (bigger jump = more energy = bluer/shorter wavelength light absorbed). Because every element and molecule has a unique set of allowed jumps, this means that the pattern of light absorbed by each element is unique - sort of a fingerprint left behind in the light.

What gets tricky is taking the amount of light absorbed (the strength of a spectral line) and using this to figure out the actual abundance of an element. The absorption you find in a line depends on both the number of absorbers present (the abundance of an element) and what fraction of that element's electrons are in the proper energy level (and able to absorb that specific color of light to jump to an allowed higher level). This will depend on conditions such as temperature, pressure (which vary with depth), and microturbulent velocity.

But if you know these quantities, you know where the electrons are. Then, by observing how much absorption takes place, you can use the strength of a spectral line to determine how many atoms of the element must be present to match the observed strength of the line (the abundance of that element in the star's atmosphere).

Temperature:
Consider the formation of spectral lines by a single element, say iron. Each of the allowed electron transitions is the result of an electron jumping upward from a specific energy level. Of course the number of electrons in each level is to first order a function of temperature (c.f., the Boltzmann equation). So, if you look at a number of different Fe lines, from a variety of different energy levels, the strengths of those lines depends on both the number of absorbers (Fe atoms) and the temperature.

Think about working the problem backwards. You know you have the same number of absorbers (Fe atoms), no matter which Fe line you look at. Now assume a certain effective temperature for the star and calculate how much Fe must be present to match the observed absorption in each of the Fe lines (i.e., looking at lines from electrons starting from many different energy levels). If you have the temperature right, you should have the same number of absorbers (same Fe abundance) for all of them.

But if you have assumed a wrong temperature, you'll find that different lines (from different initial energy levels) require different abundances to match the observed absorption. For example, if your assumed temperature is too high, you will have too many electrons in the higher energy levels and too few in the lower levels. The lines caused by electrons in the lower energy levels will be fit by larger abundances than those from the higher excitation potentials (because there are few electrons in the low energy states, more atoms are needed to match the observed absorption from such lines).

So, the test as to whether you've got the temperature right is that all the Fe observed Fe lines will require the same Fe abundance, regardless of excitation potential.

Pressure/Density/Surface Gravity
First note that these three properties are all tied together. Surface gravity is a function of mass and radius - but mass and radius are very difficult to determine, so we often just stick to surface gravity as that's what will affect the formation
of spectral lines anyway. The greater the surface gravity, the
greater the pressure and the higher the density of the gas at
any given level in the atmosphere. Note of course that
pressure and density are also related to temperature through
the perfect gas law, so to really do this, you’ll also need to
know how temperature changes with depth too (and as long
as we’re wallowing in details, you’ll also need to worry about
hydrostatic equilibrium to really figure out the pressures and
densities as a function of depth!).

But how does pressure/density affect the formation of
spectral lines? The greater the density, the larger the
number of absorbers (atoms) per cubic cm, and the stronger
the lines will be. But there’s even more to this! As discussed
above, the strength of a line also depends on the fraction of
the absorbers (atoms) with electrons that can do the
absorbing. Again temperature comes into play because we
not only have to worry about what energy level the electrons
are starting off at, but also how much ionization has taken
place.

That’s where pressure/surface gravity come into play. Sure,
the higher the temperature, greater the rate of ionization
(think about collisions). But the reverse process,
recombination, depends both on temperature and on density
(or pressure - the electron pressure to be specific) - see the
Saha Equation. The greater the density of electrons, more
frequently recombination will take place, and the fewer ions
you’ll end up with (at a given temperature).

So, when considering spectral lines, the strength of a line will
also depend on the electron pressure, which is also related
to the density and the surface gravity. As an added bonus, in
most stellar spectra you can often find lines due to atoms in
two different ionization states (metals in moderate
temperature stars are an excellent example). Again, as with
temperature, if you look at lines due to atoms in two
different ionization states, you can calculate the number of
absorbers which must be present (given the temperature
and surface gravity which determine how many of the atoms
have electrons in the proper energy level to cause the
spectral line). You’ll know you have the right surface gravity
(pressure) when you derive the same number of atoms for
the two ionization states.

Microturbulence

If you stop and think about it, you expect the gas in the
atmosphere of a star to be moving. With this motion comes
Doppler shifts, which will slightly change the wavelengths
which are absorbed (and emitted) by the atoms - some a little
too red, some a little too blue, essentially making the line
wider than it normally would be. One might first think of the
thermal motions of the atoms as the only source of line
broadening, but there are actually many of them. Others
include the natural line width due to the uncertainty principle
and pressure broadening due to nearby atoms. But one of
the most significant broadening sources in an average star is
microturbulence, a non-thermal "bulk" motion in the
atmosphere of a few km/sec.

The microturbulent motion also has an effect on spectral
lines, particularly the strong ones. This is because at the
wavelength center of a very strong line, most of the light is
absorbed (that’s why it’s so strong/dark). So, if you add more
absorbers to the mix, there’s just not that much light at that
wavelength for the new atoms to absorb, so the line really
doesn’t get any stronger. Such a line is said to be "saturated."
But if there’s also microturbulent motion, some of the
absorbers will be redshifted/blue-shifted away from the
central wavelength of the line. And there is light out there
away from the central wavelength - in the "wings" of the line
which they can absorb, increasing the total amount of light
that is absorbed in the line.

So, think about two lines, one strong, one weak. In the weak
line, if the microturbulence is low, most of the absorbers will
absorb near the central wavelength, and there are plenty of
photons around to be absorbed, so the strength of the line is
sensitive to the number of absorbers. Likewise if the
microturbulence is high - all this does is shift some absorbers
to the red/blue, but the total number of photons absorbed
(and the strength of the line) doesn’t change. In the strong
line, it’s a different story. With a low microturbulence, there
are lots of absorbers capable of absorbing near the line.
center. But with most of the photons removed, adding more absorbers doesn't change much, so the line is relatively insensitive to the abundance of absorbers. But if you increase the microturbulence, some of the atoms can absorb photons away from the line center, so you'll get more total flux removed (a stronger line) than in the case of a low microturbulence.

What all this means is that (just like above), you can predict how many atoms need to be in the atmosphere to get the absorption you see at a given temperature, surface gravity, and microturbulence. When you look at lines of different strengths due to one kind of atom, you expect to get the same required abundance regardless of strength. This will only be true if you have the correct microturbulence (assuming your temperature and gravity are correct too). If your microturbulence is too low, you'll need too many atoms for the strong lines (compared to the weak ones). If it's too high, you'll need too few.

Putting It All Together

What all the words above basically mean is that it's possible to look at the spectral lines of a star and get a pretty good idea of its temperature (effective temperature), surface gravity, and microturbulence. As an added bonus, as you figure these out, you can also figure out the abundances of different atoms.

In practice, it's a little more complicated and all these quantities are very deeply related. In fact, you have to essentially figure them out all at the same time. But if you keep in mind the following basic relations you should be able to do it:

Required abundances of a given element from each line must be the same for all excitation potentials - temperature

Required abundances of a given element must be the same for all ionization stages - surface gravity

Required abundances of a given element from each line must be the same for all equivalent widths (amount of absorption) - microturbulence

You Try It

The link below will take you to a web interface to some of our analysis codes. What you will be given is a set of measurements of line strengths in the star III-96 of the Galactic globular cluster M5. The spectra were taken on the Lick 3-m telescope and the lines were analyzed by Sneden et al. in 1992 (see their paper in the Astronomical Journal for details).

All of the measurements have been input for you. You just need to pick an effective temperature (hint: somewhere between 3750 and 8000K), surface gravity (hint: somewhere between 0.0 < log g < 5.0), and a microturbulence (hint: somewhere between 1.75 and 3.0 km/s).

Once you've picked your input parameters (and the above hints are the limits of acceptable input), the computer will calculate an appropriate model atmosphere using the MARCS program (see Gustafsson et al. 1972). This atmosphere (a temperature/pressure relation for an atmosphere) will be fed to the MOOG spectrum synthesis program (see Sneden 1973) and the abundances required to fit each of the observed lines (given your inputs) will be calculated and returned.

Look at your output. How did you do with the abundances from each line as a function of excitation potential? Pay particular attention to the elements with many lines and multiple ionization states present (Fe is a good one). Were the resulting abundances the same for Fe I as for Fe II? Did the abundance from a line depend on its equivalent width? If there are trends in these results that you don't like, try a different set of parameters.

If you have a temperature/surface gravity/microturbulence that you think works, you can also take a look at what abundances fit lines due to other elements too. The composition of a star provides valuable insight into where a star came from and what's going on in its core!

So, what is the temperature, surface gravity, and microturbulence for MS III-96?
What the fuss is all about

![Diagram showing parts of a spectral line]

Figure 2.11 The parts of a spectral line.

- **continuum** → intensity and shape (spectral energy distribution, SED) depends on temperature, radius, and so-called continuous opacity and metallicity
  - radius, temp → flux, luminosity
  - temp → radiation field, Planck curve (source function)
  - temp, pressure, metallicity → continuous opacity

- **line** → depth and width depends upon temperature, pressure, metallicity (abundances), ionization levels, excitation levels
  - temperature, pressure → line broadening
  - abundance, ionization, excitation → line depth
  - line opacity → bound-bound (transitions)
WHAT MAKES AN ABSORPTION PROFILE?

For pure absorption, solution to equation of radiative transfer is

\[ I_\lambda = I_\lambda^0 \exp(-\tau_\lambda) \]

\( \tau_\lambda \) is the optical depth at wavelength \( \lambda \)

\[ \tau_\lambda = \int \kappa_\lambda \, ds \]

integral over the path length of the line of sight, \( s \).

\( \kappa_\lambda \) is the opacity of the gas, all the gas physics and dynamics are contained in this single macro-state quantity which is a function of \( \lambda \); units are \( [\text{cm}^{-1}] \)

\[ \kappa_\lambda = n(s) \alpha_\lambda \]

\( n(s) \) = number density of absorbing particles, for example H\(^+\) atoms, \( [\text{cm}^{-3}] \)

\( \alpha_\lambda \) = absorption cross section \( [\text{cm}^2] \)

physics:

\( n(s) \) the number density of absorbing ion depends upon
- chemical composition of gas
- ionization balance of gas.

the ionization balance of gas depends upon temperature and radiation field:
- pressure of gas \( P = nkT \), which depends upon hydrodynamic state of gas

\( \star \) clearly, \( n \) can vary along the light beam path \( s \).
**WHAT MAKES AN ABSORPTION PROFILE? (cont)**

**Physics:**

- $\alpha_\lambda$, the cross section for absorption for the ion.

Cross section $\equiv \frac{\text{power removed from beam}}{\text{flux in beam}} \left[ \frac{\text{erg s}^{-1}}{\text{erg s}^{-1} \text{cm}^{-2}} \right] = \text{cm}^2$

depends upon

- atomic physics, the natural "absorption profile" for the electron transition as a function of $\lambda$; this profile is very narrow $\sim 10^{-4}\text{Å}$ centered on the transition energy $\frac{hc}{\lambda} = \Delta E_{\text{electron}}$

- line of sight velocity distribution of absorbing atoms $\rightarrow$ line of sight Doppler shifts of natural absorption profile.

these include:

- thermal motions, temperature dependence
- turbulent or microturbulent motions;
  - bulk flows, convection cells, turbulence
- collisional broadening, temperature, pressure dependence.
- rotational broadening, due to rotation of star

Final cross section is convolution of all contributions:

$$\alpha_\lambda = \alpha_{\text{nat}} \cdot \alpha_{\text{therm}} \cdot \alpha_{\text{turb}} \cdot \alpha_{\text{cell}} \cdot \alpha_{\text{rot}}$$

lots of physics in $\alpha_\lambda$!!
to get $n(s)$ you need to know gas geometry, hydrodynamic state of gas, chemical abundances of gas, ionization balance of gas, radiation field in/around gas, and temperature of gas.

To get $\alpha_\lambda$, you need temperature of gas, atomic physics of absorber transition, dynamical motions of gas, rotation rate of star.

This all requires sophisticated modeling of the gaseous environment in which the absorbing ion resides!

---

Simplifying optical depth. (Assume $\alpha_\lambda$ constant along light path)

$$\tau_\lambda = \int \alpha_\lambda ds = \int n(s)\alpha_\lambda ds = \alpha_\lambda \int n(s) ds$$

Defining the column density

$$N = \int n(s) ds \quad N \text{ in units } \text{cm}^{-2}$$

We obtain

$$\tau_\lambda = N \alpha_\lambda$$

Then

$$I_\lambda = I_\lambda^0 \exp(-N \alpha_\lambda)$$
in practice, one has free parameters for physics contained in $v_\lambda$ such as $T$, $v_{rot}$, $V_{turb}$, etc. that provide functional form of $v_\lambda$. Then, using $I_\lambda = I_\lambda^0 \exp(-Nv_\lambda)$ the absorption profile is $\chi^2$ fit for the best fit free parameters and column density $N$.

these profiles assume:

$\alpha_\lambda = \alpha_{rot} + \alpha_{therm}$

with $\sigma_v = 10$ km/s

$T = 50000$ K

note: $\frac{I_\lambda}{I_\lambda^0} = \exp(-\chi_\lambda)$

means that observed profiles provide a 1:1 mapping of $\chi_\lambda$

when $\chi \leq 1$, we call the line "optically thick"

for $\chi = 1$,

$\frac{I_\lambda}{I_\lambda^0} = e^{-1} = 0.368$

when $\chi \geq 4$, we call the line "saturated"
TWO ADDITIONAL ASPECTS OF SPECTRA.

1] the continuum shape and intensity.

2] Ionization breaks.

The continuum shape reflects the gas physics and radiative transfer at wavelengths unaffected by atomic/ion transitions.

An ionization break is treated much like an absorption line in the equation of radiative transfer, i.e., \( I_\lambda = I_\lambda^0 \exp(-N\sigma_\lambda) \). The difference is the \( \sigma_\lambda \) is smooth and continuous, and cuts off when \( \frac{hc}{\lambda} = E_{\text{ionization}} \).

Figure 4.7: Two selected absorption cross sections for neutral hydrogen in the ground state. (a) The bound-bound absorption cross section for the Ly\( \alpha \) 1215 transition, which peaks at 1215.6701 Å. (b) The bound-free absorption cross section for ionization from the ground state, which is possible only for photons having wavelengths less than \( \approx 912 \) Å.

Ionization yields sudden breaks in the continuum.
examples of an ionization edge, known as the Lyman-limit break.

- due to H\(^0\) (neutral hydrogen) atoms

\[
\tau_{\text{edge}} = -\ln\left(\frac{I_\text{obs}}{I_\lambda}\right) = -\ln(0.65) = 0.43
\]

also note convergence of higher-order neutral hydrogen transitions leading up to the break.

\[
\begin{align*}
\text{Ly} \alpha &= n=2 \rightarrow n=1 \\
\text{Ly} \beta &= n=3 \rightarrow n=1 \\
\text{Ly} \gamma &= n=4 \rightarrow n=1
\end{align*}
\]

the lines are modeled using cross section appropriate for each transition.

here we assumed \( \sigma_{vn} = 30 \text{ km/s} \).
The fundamental quantity measured in spectra is the Equivalent Width:

\[ W = \int_{\lambda_1}^{\lambda_2} \left( 1 - \frac{I_\lambda}{I_\lambda^c} \right) d\lambda \quad [\text{Å}] \]

measures span of wavelength for an "ideal" square black bottom absorption feature.

**Figure 16.4:** A schematic of four absorption lines each with equivalent width \( W = 0.21 \text{ Å}. \) Though the line profile shapes are quite different, the total amount of flux absorbed, as given by Eq. 16.15, are identical. The shaded grey area shows the interpretation of the equivalent width.

As we showed, \( \tau_\lambda = N \alpha_\lambda \), \( \frac{I_\lambda}{I_\lambda^c} = \exp(-N \alpha_\lambda) \)

Thus

\[ W = \int_{\lambda_1}^{\lambda_2} \left( 1 - \exp(-N \alpha_\lambda) \right) d\lambda \]

\( \alpha_\lambda \) contains \( \sigma_v \) info (velocity dispersion of absorbing atoms) define \( b = \sqrt{2} \cdot \sigma_v \) so that \( \alpha_\lambda = \alpha_\lambda(b) \) \{and other gos parameters\}
Figure 16.5: The COG showing equivalent width as a function of the optical depth at line core, \( \tau_c \), for Ly\\alpha for \( b = 30 \) km sec\(^{-1} \). The three regimes, "linear", "logarithmic", and "square root", corresponding to Eq. 16.17, are shown by the thick curves, respectively, as \( \tau_c \) increases. Absorption profiles are shown for each regime and their locations on the COG marked with filled points. Note the expanded wavelength scale for the profiles on the square root part of the COG. This is due to large damping wings.

Figure 16.6: The curve of growth for Ly\\alpha \( \lambda 1215 \), Mg\\II \( \lambda 2796 \), CIV \( \lambda 1548 \), and OVI \( \lambda 1031 \). For each species, the equivalent width, \( W, [\AA] \) is plotted vs. the logarithm of the column density. For the curves shown, the range of \( b \) parameters are given in the lower right corner of the panels, with \( b \) increasing in steps of \( \Delta b \) upward.
The Photosphere

Where the Spectrum Comes From

Photosphere is where escape probability is near unity -
- quantified by optical depth
\[ \tau = \int n_\text{gas} \, ds \]
- for absorption
\[ I_\lambda = I_0 \exp \left( -\tau_\lambda \right) \]

Photosphere lines arise in region \( \frac{2}{3} \leq \tau \leq 2 \) (roughly)

- For the sun, Fig 1.1 shows \( T \) vs height where \( h = 0 \, \text{km} \) is at \( \tau = 1 \)
  - note narrow region of line formation
- Fig "1.2" (to the right) show \( T \) and \( n_e \) from the "solar surface" (\( h = 0 \)) in units of solar radii (top) and in cm (top).

Fig. 1.1. The temperature distribution in the outer layers of the Sun is shown as a function of the geometrical depth in kilometers. The center of the Sun is toward the right. The photosphere is indicated by the shaded region. The visible-light spectrum comes from the photosphere. Adapted from Vernazza et al. (1973).


\[ n_e \approx 10^{13-14} \, \text{cm}^{-3} \] in solar photosphere
- decreasing 3-4 orders of magnitude by the time the chromosphere is reached.

\[ n_{\text{tot}} \approx 10^{17} \, \text{cm}^{-3} \] (see next page)
Figure 2.1 The OBAFGKM spectral sequence for main-sequence (dwarf) stars illustrating that the spectral sequence is ordered in terms of temperature. Here, the normalized stellar flux (the energy distribution) is plotted against wavelength. Some of the more prominent spectral features are marked, including the Balmer jump and convergence. The source of these spectra is the Indo-US coude-feed spectral library (Valdes et al. 2004). The spectra have been normalized at a common wavelength, and separated by one continuum unit for clarity, except for the bottom spectrum, which is offset by 2 units.
Hydrogen Series

Table 11. The wavelengths in Å of the $m \rightarrow n$ transitions of hydrogen for $n = 1$ to 6, $m = 2$ to 21, and $m = \infty$, and for the $n = 4$ Pickering series for ionized helium (He II)\(^1\). Here the wavelengths are in Å where 1 Å = 10\(^{-8}\) cm.

<table>
<thead>
<tr>
<th>Series</th>
<th>Lyman ($n=1$)</th>
<th>Balmer ($n=2$)</th>
<th>Paschen ($n=3$)</th>
<th>Bracket ($n=4$)</th>
<th>Pfund ($n=5$)</th>
<th>Humphreys ($n=6$)</th>
<th>Pickering (He (^+), $n=4$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>2</td>
<td>1,215.67 Å</td>
<td>6,562.8 Å</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>3</td>
<td>1,025.72 Å</td>
<td>6,604.5 Å</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>4</td>
<td>972.53 Å</td>
<td>4,861.3 Å</td>
<td>18,751 Å</td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>5</td>
<td>949.74 Å</td>
<td>4,340.4 Å</td>
<td>12,818.1 Å</td>
<td>40,542.0 Å</td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>6</td>
<td>937.80 Å</td>
<td>4,101.7 Å</td>
<td>10,938.1 Å</td>
<td>26,252.0 Å</td>
<td>74,578 Å</td>
<td>10,123.6 Å</td>
<td></td>
</tr>
<tr>
<td>7</td>
<td>930.74 Å</td>
<td>3,970.0 Å</td>
<td>10,049.4 Å</td>
<td>21,655.0 Å</td>
<td>46,525 Å</td>
<td>123,680 Å</td>
<td>6,560.10 Å</td>
</tr>
<tr>
<td>8</td>
<td>926.22 Å</td>
<td>3,899.0 Å</td>
<td>9,545.9 Å</td>
<td>19,445.6 Å</td>
<td>37,395 Å</td>
<td>75,005 Å</td>
<td>4,859.32 Å</td>
</tr>
<tr>
<td>9</td>
<td>923.15 Å</td>
<td>3,835.3 Å</td>
<td>9,229.0 Å</td>
<td>18,174.1 Å</td>
<td>32,961 Å</td>
<td>59,066 Å</td>
<td>4,541.59 Å</td>
</tr>
<tr>
<td>10</td>
<td>920.96 Å</td>
<td>3,797.9 Å</td>
<td>9,014.9 Å</td>
<td>17,362.1 Å</td>
<td>30,384 Å</td>
<td>51,273 Å</td>
<td>4,338.67 Å</td>
</tr>
<tr>
<td>11</td>
<td>919.35 Å</td>
<td>3,770.6 Å</td>
<td>8,862.7 Å</td>
<td>16,806.5 Å</td>
<td>28,722 Å</td>
<td>46,712 Å</td>
<td>4,199.83 Å</td>
</tr>
<tr>
<td>12</td>
<td>918.12 Å</td>
<td>3,750.1 Å</td>
<td>8,750.4 Å</td>
<td>16,407.2 Å</td>
<td>27,575 Å</td>
<td>43,753 Å</td>
<td>4,100.04 Å</td>
</tr>
<tr>
<td>13</td>
<td>917.18 Å</td>
<td>3,734.3 Å</td>
<td>8,665.0 Å</td>
<td>16,109.3 Å</td>
<td>26,744 Å</td>
<td>41,697 Å</td>
<td>4,025.60 Å</td>
</tr>
<tr>
<td>14</td>
<td>916.42 Å</td>
<td>3,721.9 Å</td>
<td>8,598.3 Å</td>
<td>15,880.5 Å</td>
<td>26,119 Å</td>
<td>40,198 Å</td>
<td>3,968.43 Å</td>
</tr>
<tr>
<td>15</td>
<td>915.82 Å</td>
<td>3,711.9 Å</td>
<td>8,545.3 Å</td>
<td>15,700.7 Å</td>
<td>25,636 Å</td>
<td>39,065 Å</td>
<td>3,923.48 Å</td>
</tr>
<tr>
<td>16</td>
<td>915.32 Å</td>
<td>3,703.8 Å</td>
<td>8,502.4 Å</td>
<td>15,556.5 Å</td>
<td>25,254 Å</td>
<td>38,184 Å</td>
<td>3,887.44 Å</td>
</tr>
<tr>
<td>17</td>
<td>914.91 Å</td>
<td>3,697.1 Å</td>
<td>8,467.2 Å</td>
<td>15,438.9 Å</td>
<td>24,946 Å</td>
<td>37,484 Å</td>
<td>3,858.07 Å</td>
</tr>
<tr>
<td>18</td>
<td>914.57 Å</td>
<td>3,691.5 Å</td>
<td>8,437.9 Å</td>
<td>15,341.8 Å</td>
<td>24,693 Å</td>
<td>36,916 Å</td>
<td>3,833.80 Å</td>
</tr>
<tr>
<td>19</td>
<td>914.28 Å</td>
<td>3,686.8 Å</td>
<td>8,413.3 Å</td>
<td>15,260.6 Å</td>
<td>24,483 Å</td>
<td>36,449 Å</td>
<td>3,813.50 Å</td>
</tr>
<tr>
<td>20</td>
<td>914.03 Å</td>
<td>3,682.8 Å</td>
<td>8,392.4 Å</td>
<td>15,191.8 Å</td>
<td>24,307 Å</td>
<td>36,060 Å</td>
<td>3,796.33 Å</td>
</tr>
<tr>
<td>21</td>
<td>913.82 Å</td>
<td>3,679.3 Å</td>
<td>8,373.5 Å</td>
<td>15,127.0 Å</td>
<td>24,154 Å</td>
<td>35,741 Å</td>
<td>3,781.68 Å</td>
</tr>
<tr>
<td>∞</td>
<td>911.5 Å</td>
<td>3,646.0 Å</td>
<td>8,203.6 Å</td>
<td>14,584 Å</td>
<td>22,788 Å</td>
<td>32,814 Å</td>
<td>3,644.67 Å</td>
</tr>
</tbody>
</table>

\(^1\) Data from Whish, Smith, and Glennon (1966).
Hydrogen Lines and Spectral Regions

*Lyman Jump* → *Balmer Jump* → *Paschen Jump* → *Brackett Jump*

![Diagram of Hydrogen Spectral Series](image)

**Fig. 3-3** Spectral Series in Hydrogen between 0 and 15,000 Å. The Lyman and Balmer series are complete in this figure, but Paschen α (18,750 Å) lies to the right of the diagram. For the Brackett series, only the lines near the limit have wavelengths within the range of the figure. (Narrowing)

- **Lyman series (UV)**: upward (absorption) transitions from $n=1$ to the ground state, $T \leq 10,000$.
- **Balmer series (Visible)**: upward transitions from $n=2$; hydrogen must be in 1st excited state, $T \sim 10,000$.
- **Paschen series (IR)**: upward transitions from $n=3$; hydrogen must be in 2nd excited state, $T > 10,000$.
- **Continuum regions are named by excitation level for which ionization occurs, not bound-bound transitions**.
- **Paschen continuum** in visible, where ionization from $n=3$ occurs and Balmer lines appear.