Stellar Populations

Physics of Galaxies 2012

part 3

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Why do we need to know about stars to study galaxies?

- We can’t see the *vast majority of the mass* in galaxies
  - dark matter!
- We only see the “baryons” in galaxies
  - stars
  - gas
- To understand the properties of galaxies, we need to study these components

Stellar atmospheres

- When we look at galaxies in the UV, visible, and NIR, we see (mostly) the atmospheres of their stars
    - their continua: blackbody curves
    - their absorption lines
        - ...and sometimes the emission lines of the gas, if there is star formation or an active galactic nucleus

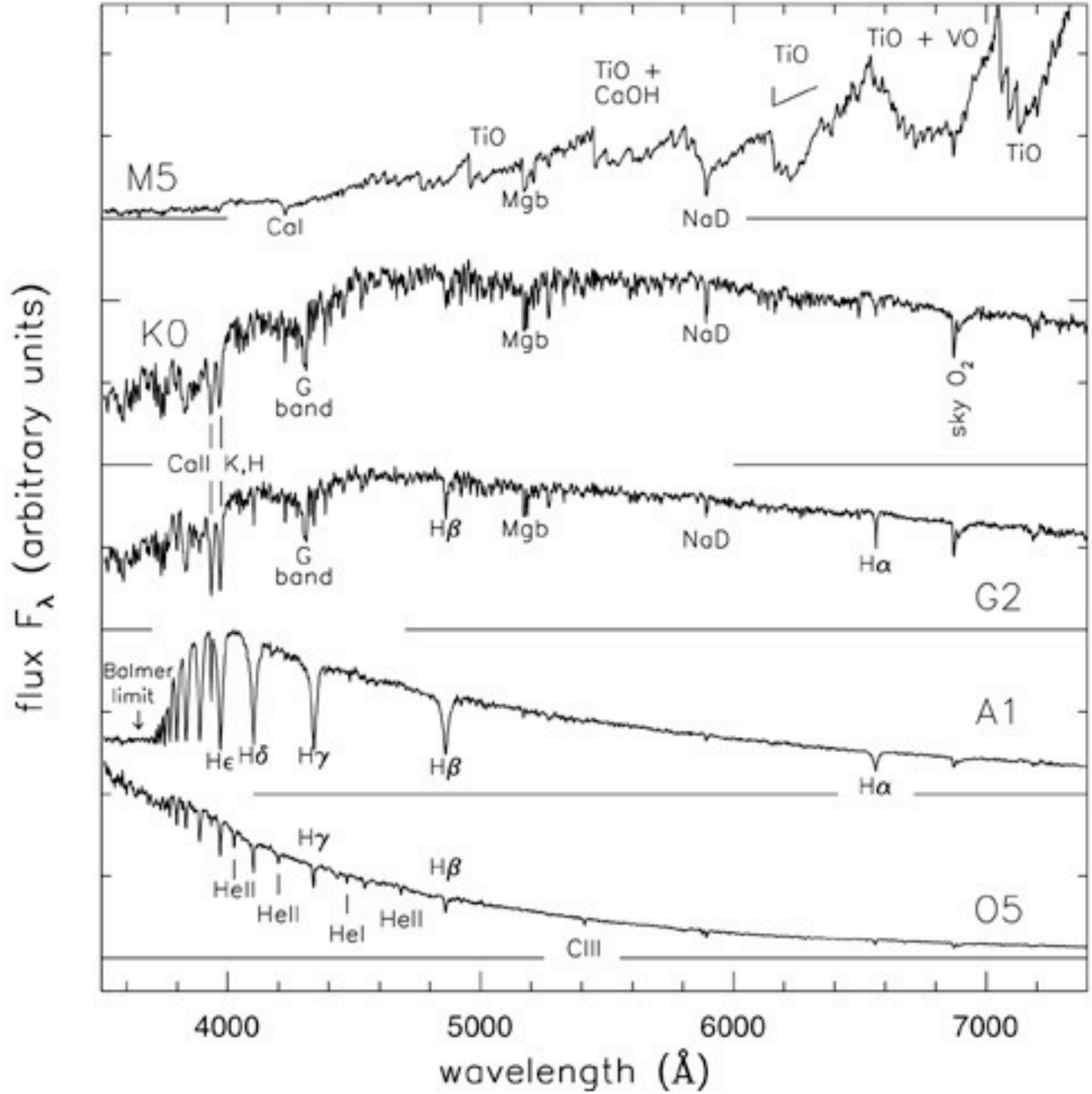
- Stellar spectra depend mainly on **three variables**:
  - *Effective temperature*, where $L \equiv 4\pi R^2 \sigma T_{\rm eff}^4$
  - To first order, the emitted continuum of a star is a blackbody curve with $T \sim T_{\rm eff}$
    - and a peak at $\lambda_{\rm max} = [2.9/T\,({\rm K})]\,{\rm mm}$

- *Composition*: more-or-less the mean abundance of elements heavier than He

$$[\mathrm{Fe/H}] \equiv \log_{10}[\epsilon(\mathrm{Fe/H})_*] - \log_{10}[\epsilon(\mathrm{Fe/H})_\odot]$$

where **ε**(Fe/H) is the number of Fe atoms relative to the number of H atoms

  - causes absorption lines where atoms "intercept" continuum light of star
    - also "line blanketing": when absorption lines "pile up", like below 4000 Å (not to be confused with Balmer decrement...)


M5
TiO
TiO +
CaOH
TiO
TiO + VO
TiO
Mgb
CaI
NaD
K0
Mgb
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G
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CaII K,H
Hβ
Mgb
G
band
NaD
Hα
G2
Balmer
limit
A1
Hε Hδ
Hγ
Hβ
Hα
Hγ
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HeII
HeII
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HeII
O5
CIII
4000
5000
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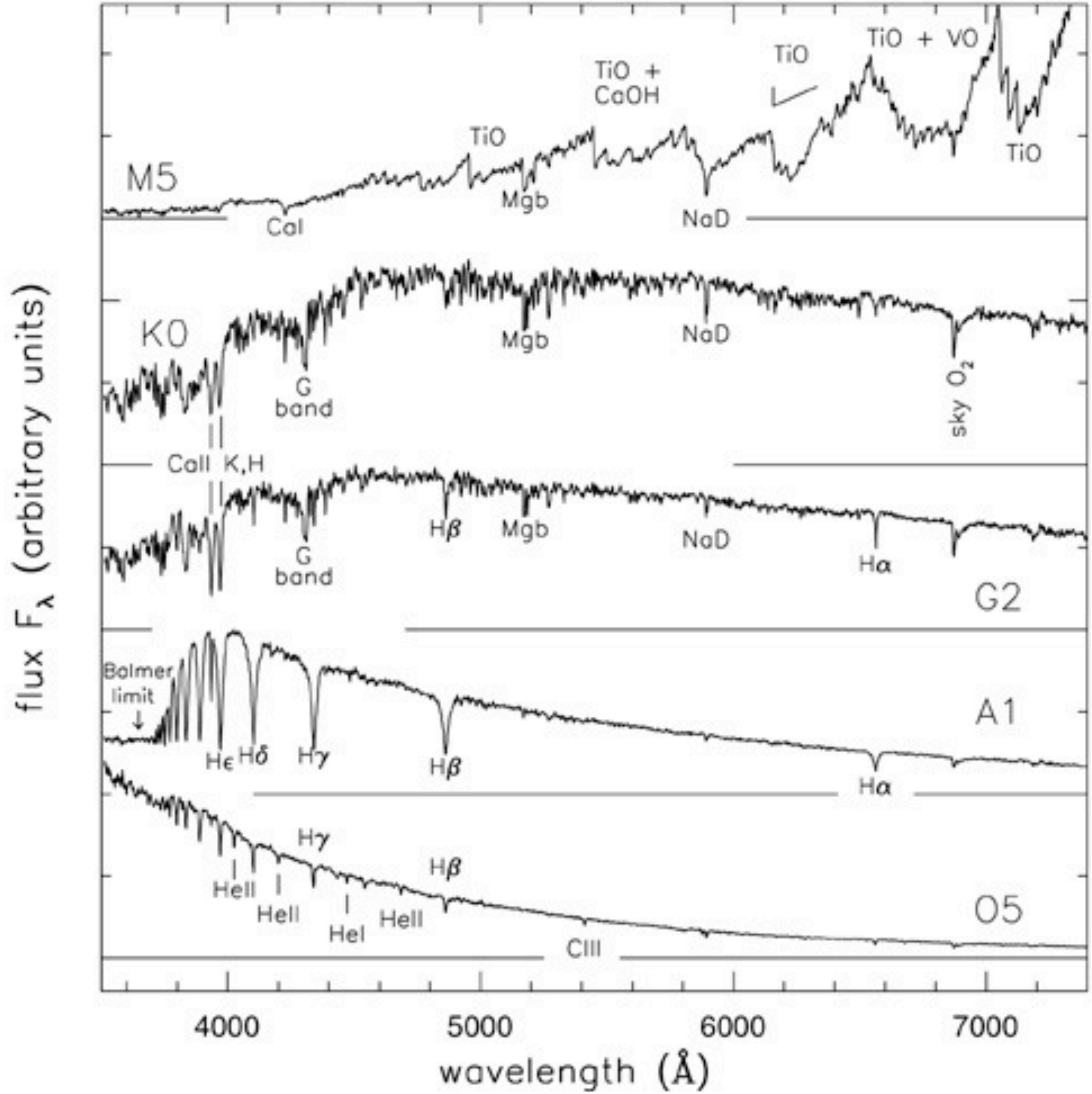
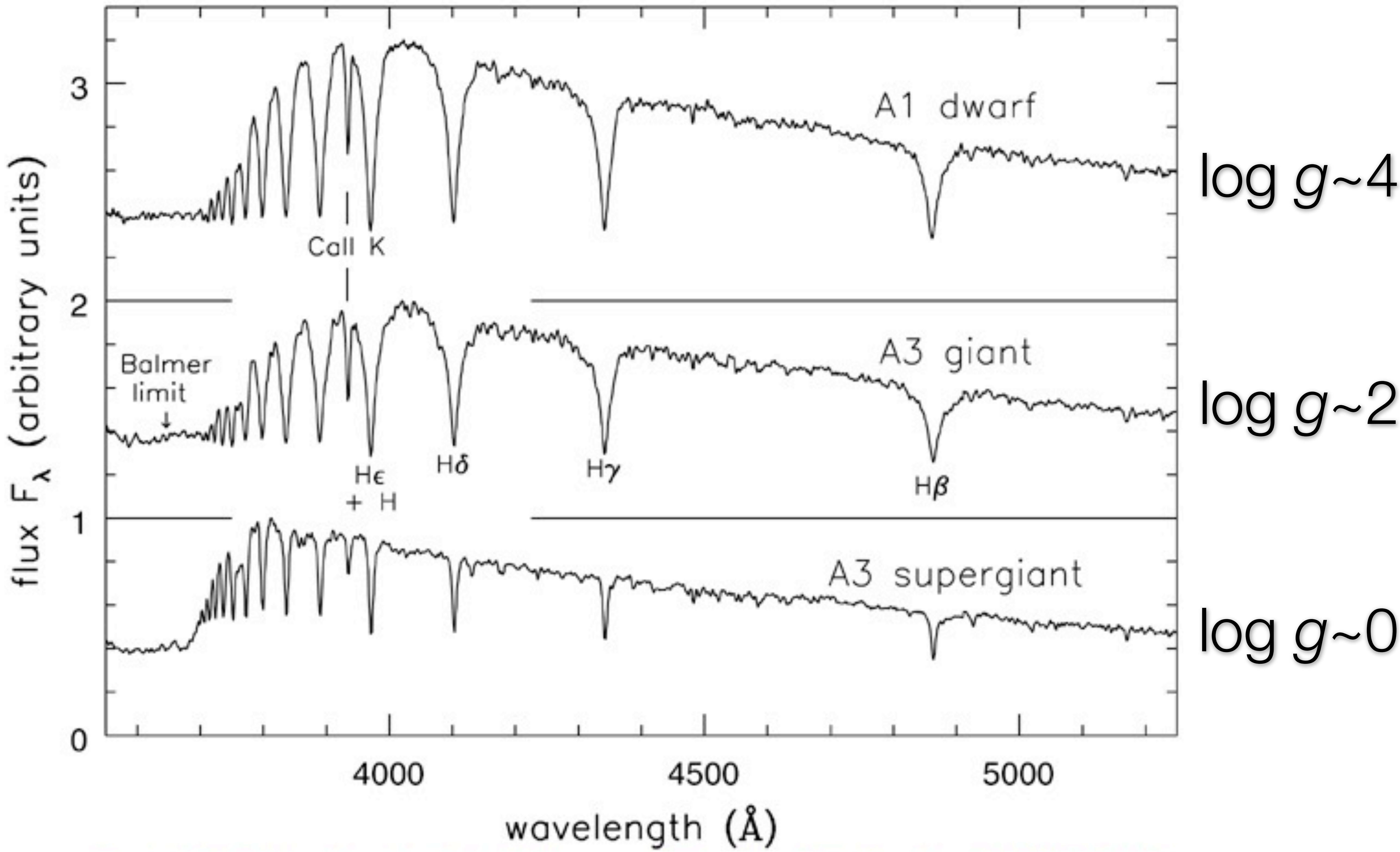
7000
wavelength (Å)
flux Fλ (arbitrary units)



Cooler

Hotter

- *Surface gravity*: $g \equiv \dfrac{GM}{R^2}$
  - set gas pressure
    - ionization balance (e.g., singly- vs. doubly-ionized Fe)
    - pressure broadened lines (like H lines)
    - abundances of certain molecules (in cool stars: CN, CO, etc.)



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Stellar classification (quickly)

- Morgan-Keenan spectral classes
  - based on ratios of strong absorption lines
  - OBAFGKM(LT) ⇒ T_{eff}
  - Luminosity class I-V ⇒ log *g*
    - Sun is G2V star
  - Only works within ~2x of solar abundance; qualitative; discrete steps; no abundance calibration

Magnitudes and colors

- Unfortunately, spectra are *expensive* in telescope time
  - need to spread photons out into lots of pixels – takes a lot of time to do!
- Need faster way to get stellar properties

- Consider two objects with fluxes f_1 and f_2
- Then the *magnitude difference* between these objects is $$m_1 - m_2 = -k \log_{10}\left(\frac{f_1}{f_2}\right)$$
  - where $k = 2.5$ means that $m_1 - m_2 = 5$ when $f_2/f_1 = 100$
  - so $m_1 - m_2 = -2.5 \log_{10}\left(\frac{f_1}{f_2}\right)$
  - or $\dfrac{f_1}{f_2} = 10^{-0.4(m_1 - m_2)}$

note: from here on, I mean log-base-10 for "log" and log-base-e for "ln"

- This defines the **apparent magnitude**, the magnitude of the flux received by the detector at the telescope, where

$$f \equiv \int_0^\infty f_\nu F_\nu R_\nu T_\nu d\nu$$

  - here
    - f_ν is the flux of the object (in frequency units)
    - F_ν is the transmission of any filter used to isolate the (frequency) region of interest
    - R_ν is the transmission of the telescope, optics, and detector
    - T_ν is the transmission of the atmosphere (if any)

- Most (but not all) magnitude systems are based on taking a magnitude with respect to a star with a known (or predefined) magnitude
  - the “Vega” system defines a set of A0V stars as having apparent magnitude 0 in *all* bands of a system (see below)
  - So to get a “magnitude on system X”, one observes stars with known magnitudes and *calibrates* the “instrumental magnitudes” onto the “standard system”

- Some useful properties and "factoids" about magnitudes...
  - The magnitude system is roughly based on *natural logarithms*: $m_1 - m_2 = 0.921 \ln(f_1/f_2)$
  - If $\Delta f \ll 1$, then $\Delta m = m_2 - m_1 \approx 1.086 \Delta f$
    - so the magnitude difference between two objects of nearly-equal brightness is equal to the fractional difference in their brightnesses – i.e., a difference of 0.1 magnitudes is ~10% in brightness
  - A factor of 2 difference in brightness is a difference of 0.75 magnitudes

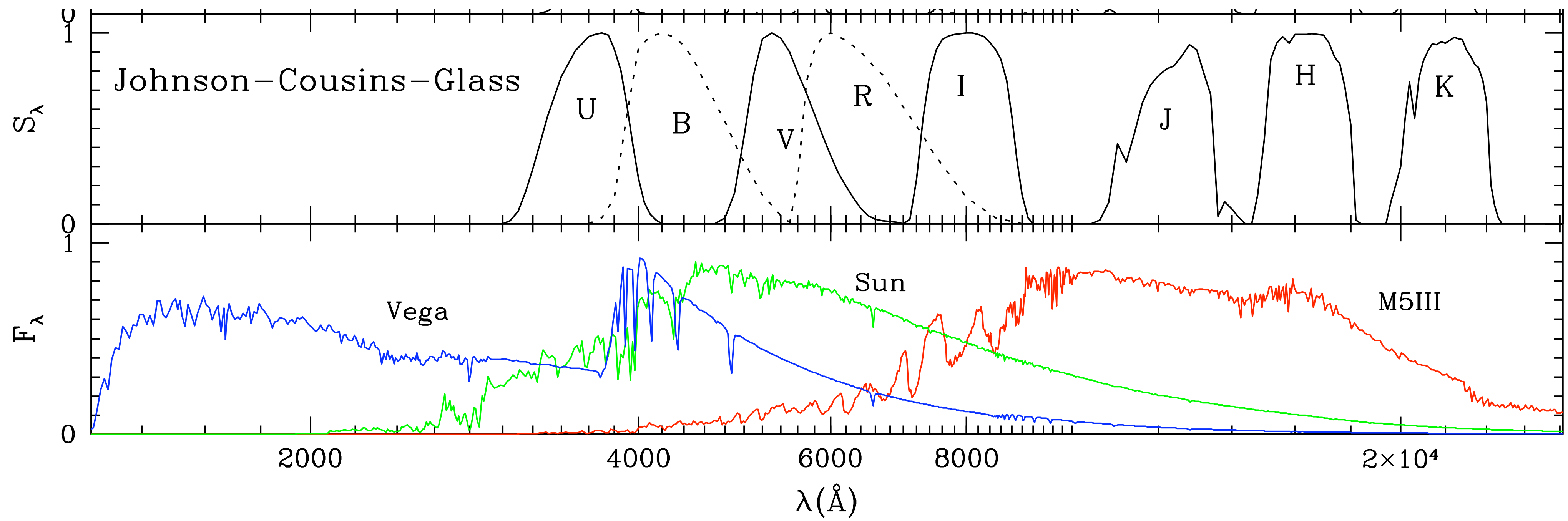
- We define the **color** of an object as the magnitude difference of the object in two different filters ("bandpasses")

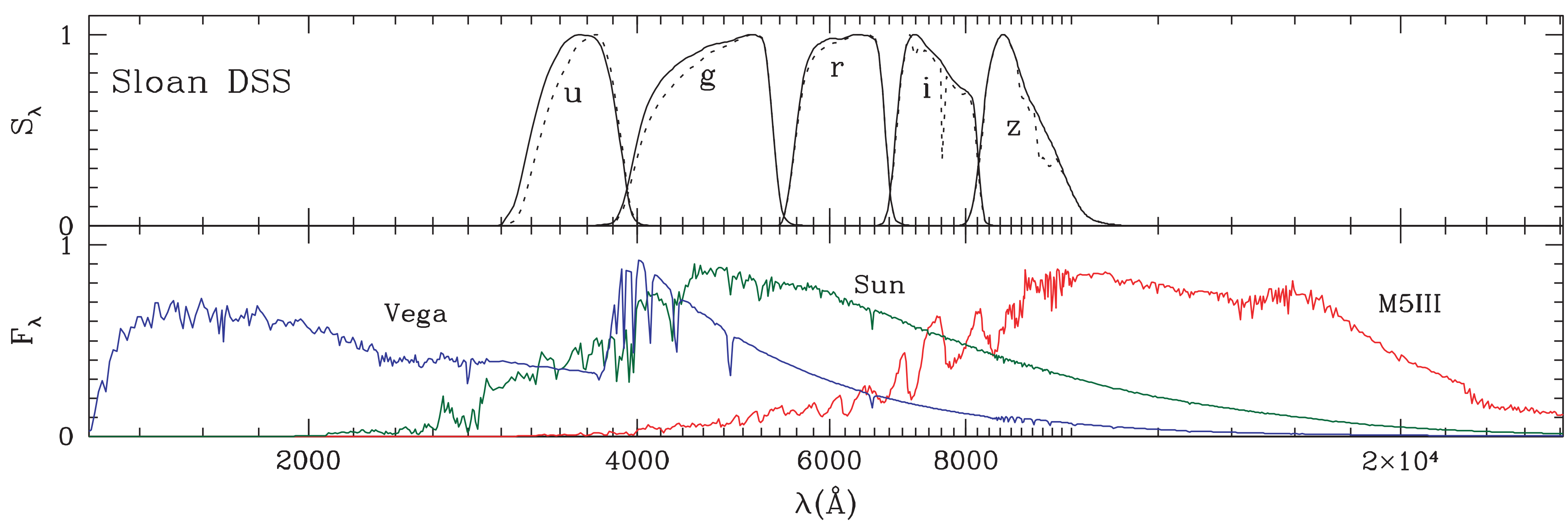
  - if the filters are *X* and *Y*, then the color (*X*–*Y*) is

$$
\begin{aligned}
(X-Y) \equiv m_X - m_Y &= \text{const.} - 2.5 \log \frac{\int_0^\infty d\lambda S_\lambda(X) f_\lambda}{\int_0^\infty d\lambda S_\lambda(Y) f_\lambda} \\
&= \text{const.} - 2.5 \log \frac{f_X}{f_Y}
\end{aligned}
$$

    - where S_λ is the combined telescope-detector-filter sensitivity

Two common magnitude systems





- Notice that stars of different temperatures have very different spectra, so they have different fluxes through the filters
  - therefore they have *different colors*
  - and therefore well-chosen colors can be good **temperature indicators**
    - problems: calibration is tricky – requires good stellar radii; line blanketing makes bands bluer than ~V sensitive to [Fe/H] and log g

Apparent to absolute magnitudes

strictly true only if M is constant, but this is almost true on the giant branch.

- At fixed effective temperature, stars with lower log *g* must be bigger, since $g \equiv \frac{GM}{R^2}$
  - so they must have higher luminosity, since $L \equiv 4\pi R^2 \sigma T_{\mathrm{eff}}^4$
- But we measure the flux, not the luminosity: $F = \dfrac{L}{4\pi D^2}$
- so we need the *distance D to the stars*

- We will come back to how to determine distances in the next lecture. In the meantime...

- We receive flux *f* at our telescopes from an object at distance *d*. If we wanted instead to know what the flux *F* is of that object if it were at distance *D*, then

$$f = \left(\frac{D}{d}\right)^2 F$$

- The difference in magnitudes *m* of *f* at *d* and *M* of *F* at *D* is then

$$m - M = -2.5 \log\left(\frac{f}{F}\right) = 5 \log\left(\frac{d}{D}\right)$$

- Now, we *always* pick *D* to be 10 parsecs (10 pc), so if *d* is measured in parsecs, then

$$m - M = 5\log(d[\mathrm{pc}]) - 5$$

- If *d* is measured in 10^6 pc=1 Mpc, then

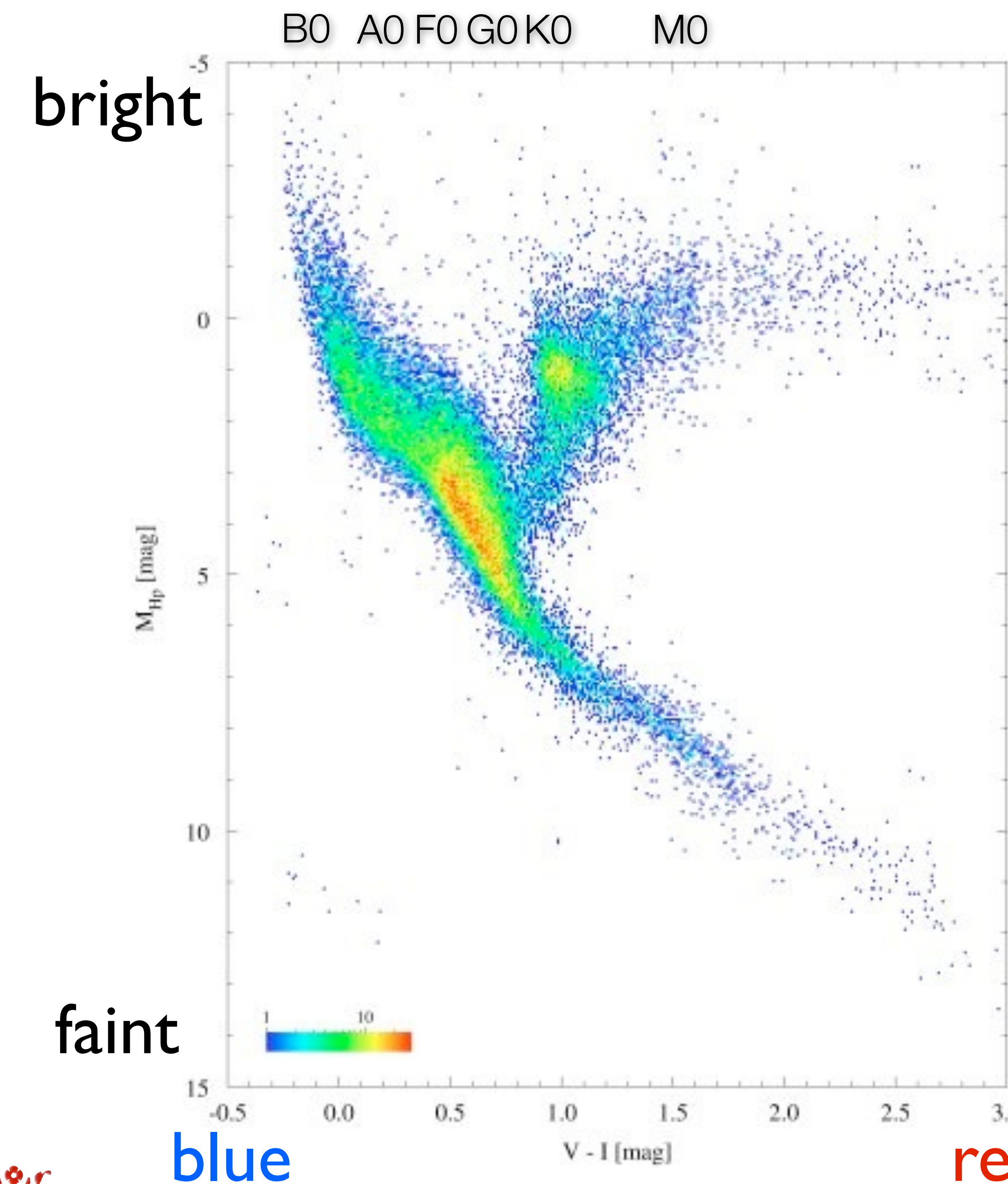
$$m - M = 5\log(d[\mathrm{Mpc}]) + 25$$

- *m–M* is called the **distance modulus** and is sometimes written μ=*m–M*

- ...and *M* is called the **absolute magnitude**

Color-magnitude diagrams

- By combining precisely-measured distances with apparent magnitudes and colors, we can plot a **color-magnitude diagram** (CMD) of stars
    - This CMD is for stars within 100 pc of the Sun with accurate distances and colors



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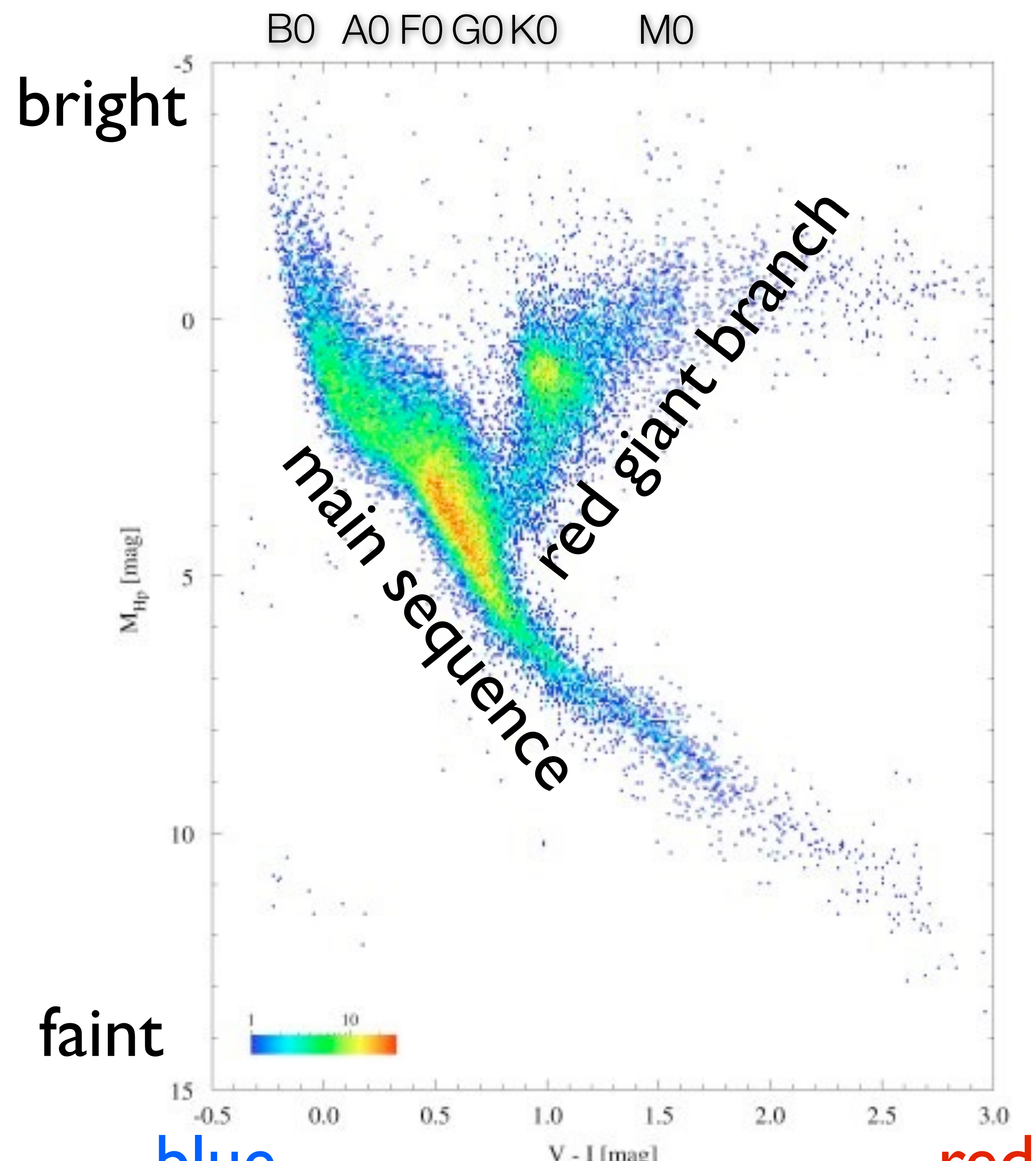
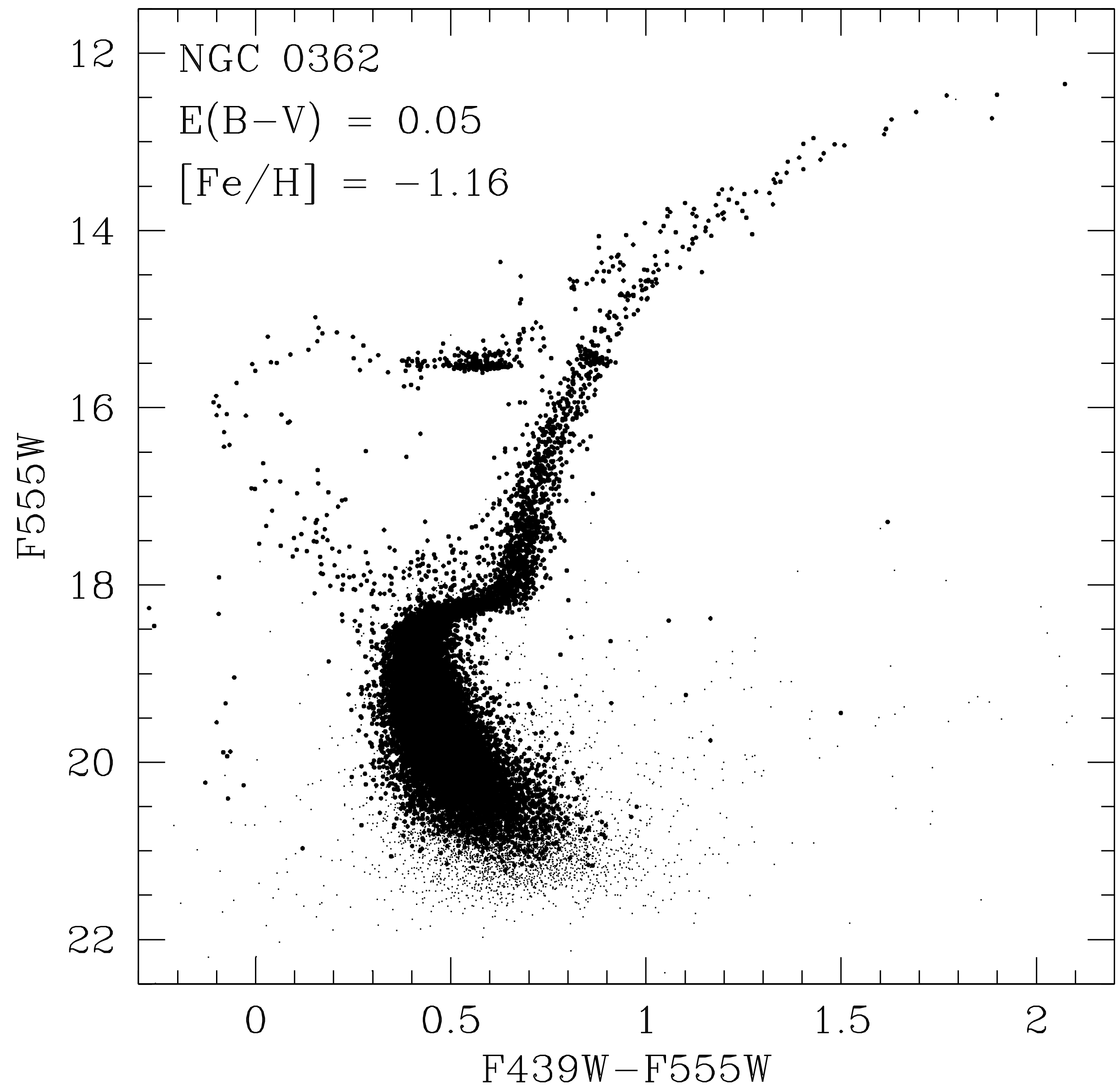


- Note!
  - A **Hertzsprung-Russell diagram** is a plot of luminosity as a function of effective temperature
  - A **color-magnitude diagram** is a plot of magnitude as a function of color
  - *They are not the same!*
    - The conversion from color to temperature depends on knowing the *radii* of stars, which (as mentioned) are difficult to measure!
    - Magnitude $\neq$ luminosity unless you know the distance and the *bolometric correction*, which depends on stellar type (or color)

- The local CMD is made up of stars of many ages (we'll return to this point!)
- When we look at an object like a globular cluster, we see a very distinct pattern



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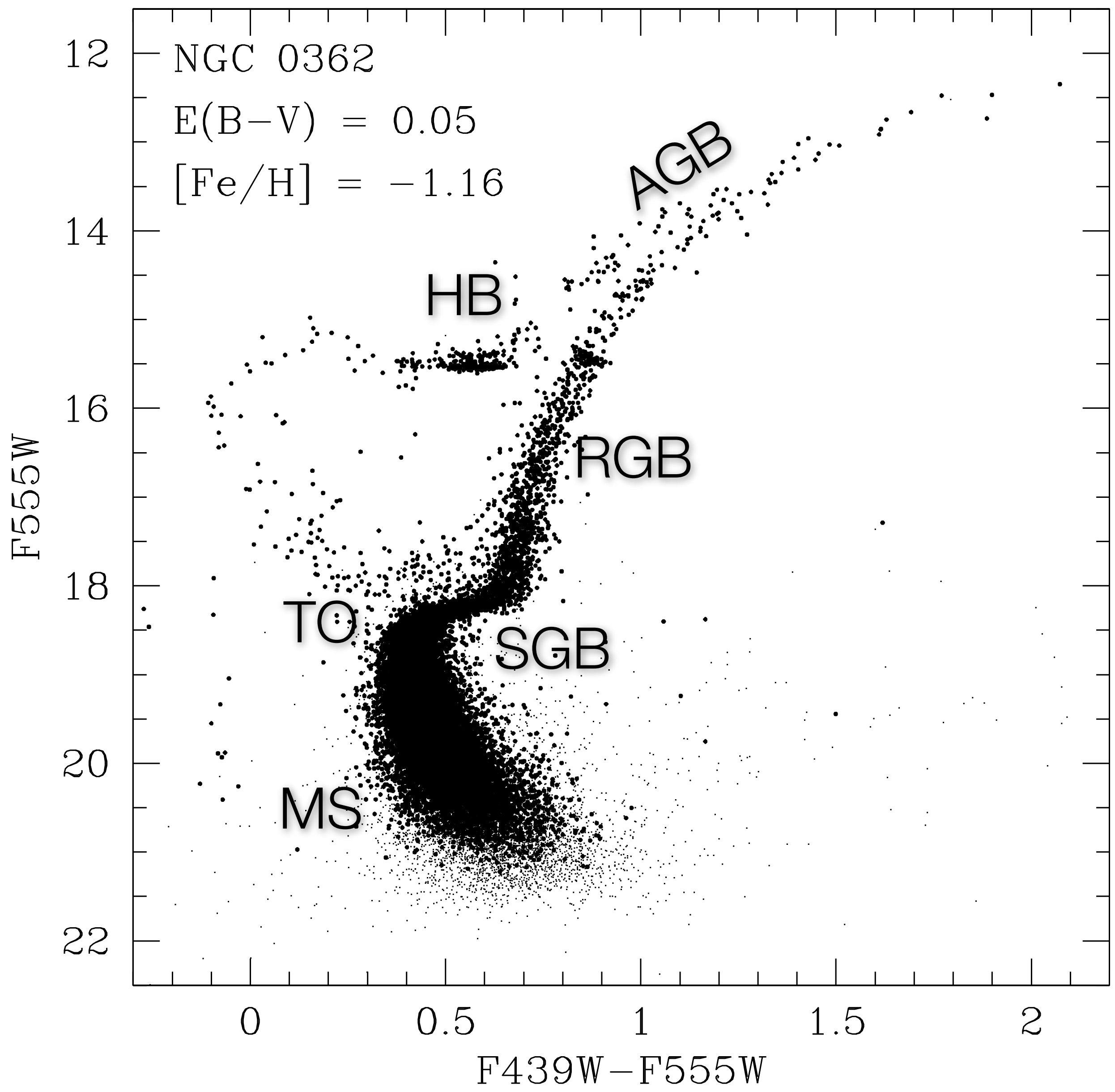

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Evolutionary timescales

- Stars burn hydrogen into helium for most of their lifetimes, until they exhaust the H in their cores
  - How long does this take?
    - Let's call *E* the amount of energy released by H burning during its *main-sequence lifetime* $\tau_{\rm MS}$
    - If the star's luminosity on the main sequence is *L*, then $E = L\tau_{\rm MS}$
    - So if we can determine *E*, we can determine $\tau_{\rm MS}$

- Now, the amount of energy released by converting a mass *dM* of H into He is $dE = 0.007c^2 dM$
- So if a fraction **α** of the total mass of the star *M* is burned into He before core H-exhaustion, then $E = 0.007c^2 \alpha M$
- so the main-sequence lifetime is $\tau_{\rm MS} = 0.007 \alpha M c^2 / L$
- Typically, **α**=0.1, so $\tau_{\rm MS} = 10 M/M_\odot (L/L_\odot)^{-1}\,{\rm Gyr}$

- It is straightforward to work out from the equations of stellar structure that $L \propto M^4$ over a good chunk of the main sequence (for stars near the solar composition; this is also true observationally!)

- Then finally we have the main-sequence lifetime as a function of mass: $\tau_{\mathrm{MS}} = 10A(M/M_{\odot})^{-3}\,\mathrm{Gyr}$

  - where *A* is a constant of ~1

- So a star ten times the mass of the sun lives for 1/1000 of the time: 10 Myr!

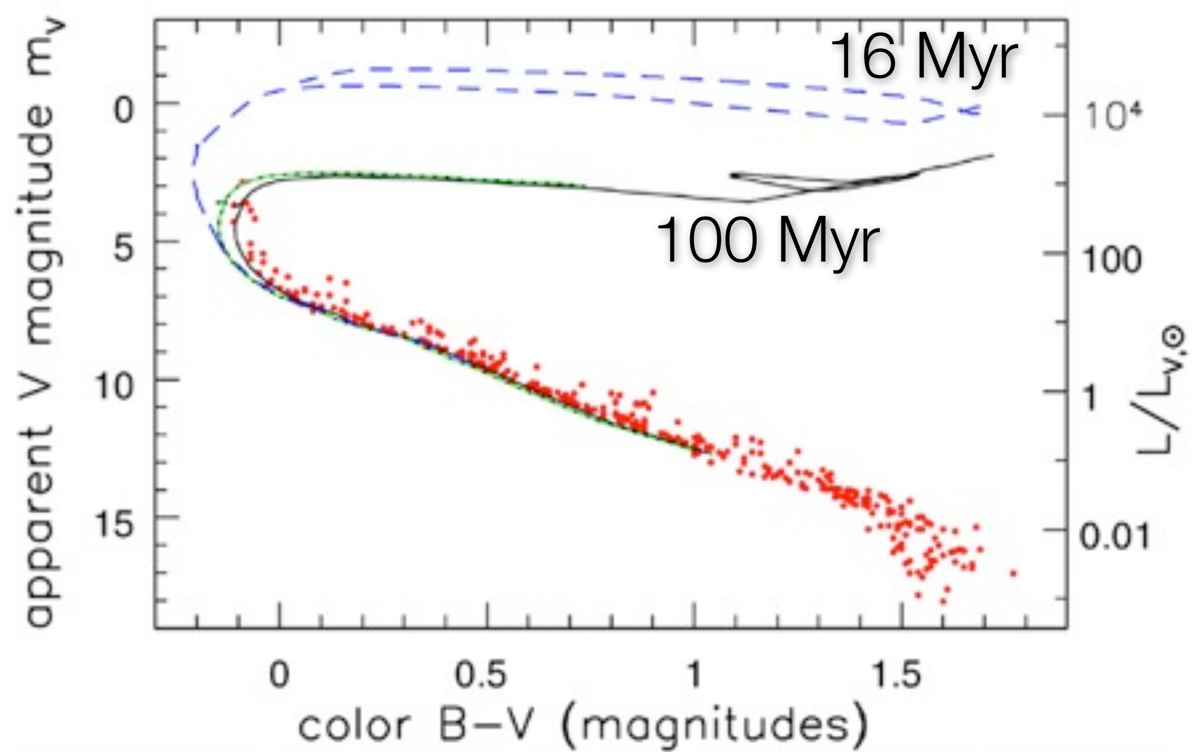
- The main-sequence lifetime is a crucial timescale: it is a *clock*
  - By measuring the magnitude and color of the **main sequence turnoff** (MSTO or just TO) we can determine the age of a stellar population – and therefore something about its evolution
  - The brighter and bluer the MSTO, the younger the population
    - brighter, bluer MS stars are more massive

Open or Galactic Clusters

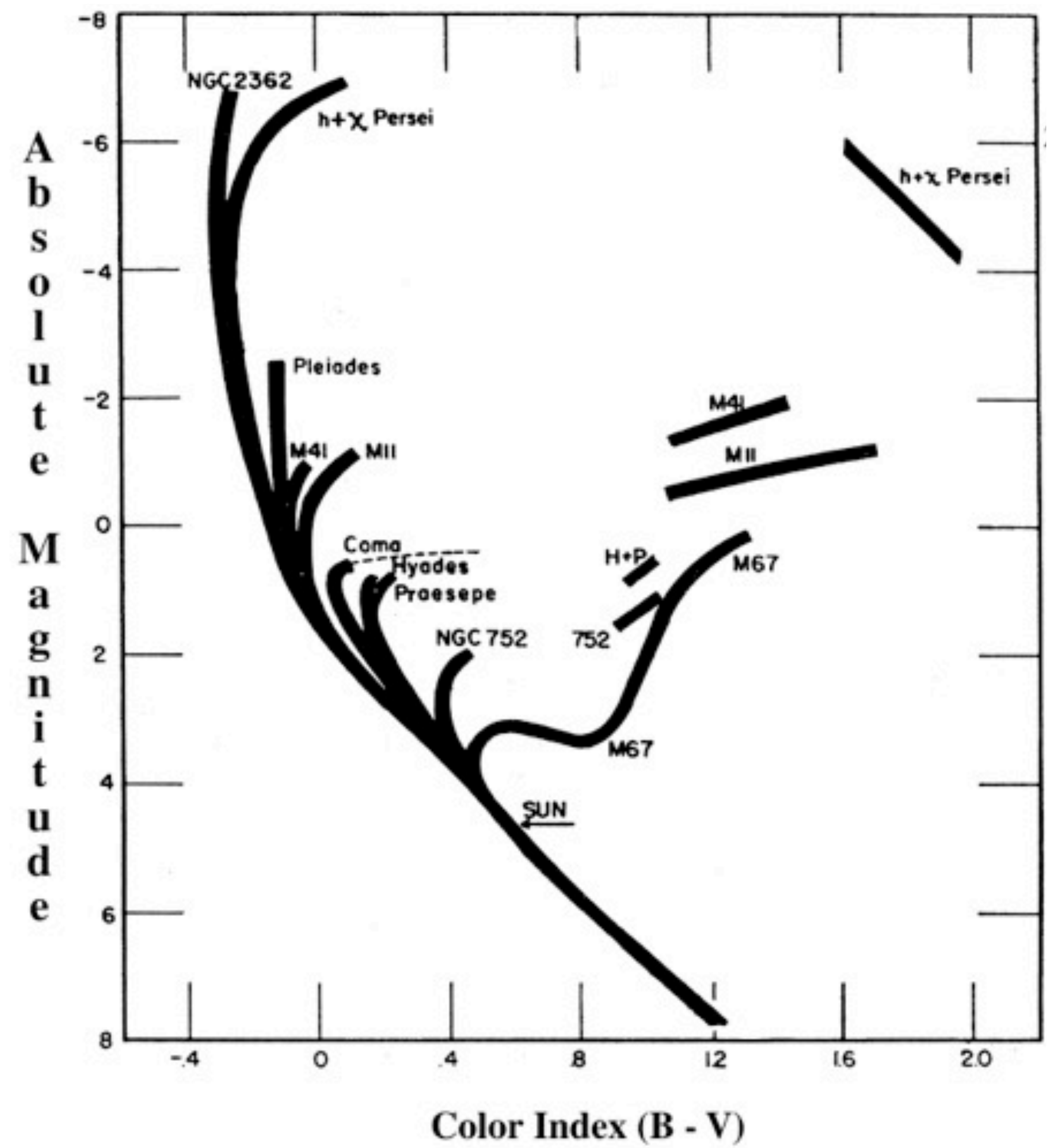
- “Open” or Galactic clusters are low mass, relatively small (~10 pc diameter) clusters of stars in the Galactic disk containing $<10^3$ stars
  - The Pleiades cluster is a good example of an open cluster
    - the “fuzziness” is starlight reflected from interstellar dust

- Because open clusters live in the disk of our Milky Way, they are subject to strong tides and shearing motions
- Because they are so small and contain few stars, they also *evaporate* quickly
- Therefore they do not live very long unless they are very massive — so most of them are quite *young*

- All stars in an open cluster
  - are at the same distance
  - formed at the same time
  - have the same composition
- Very useful for testing stellar evolution models!



- They are also very useful for studying the evolution of the properties of the MW's disk
  - Which of these clusters is the youngest?
  - Which is the oldest?



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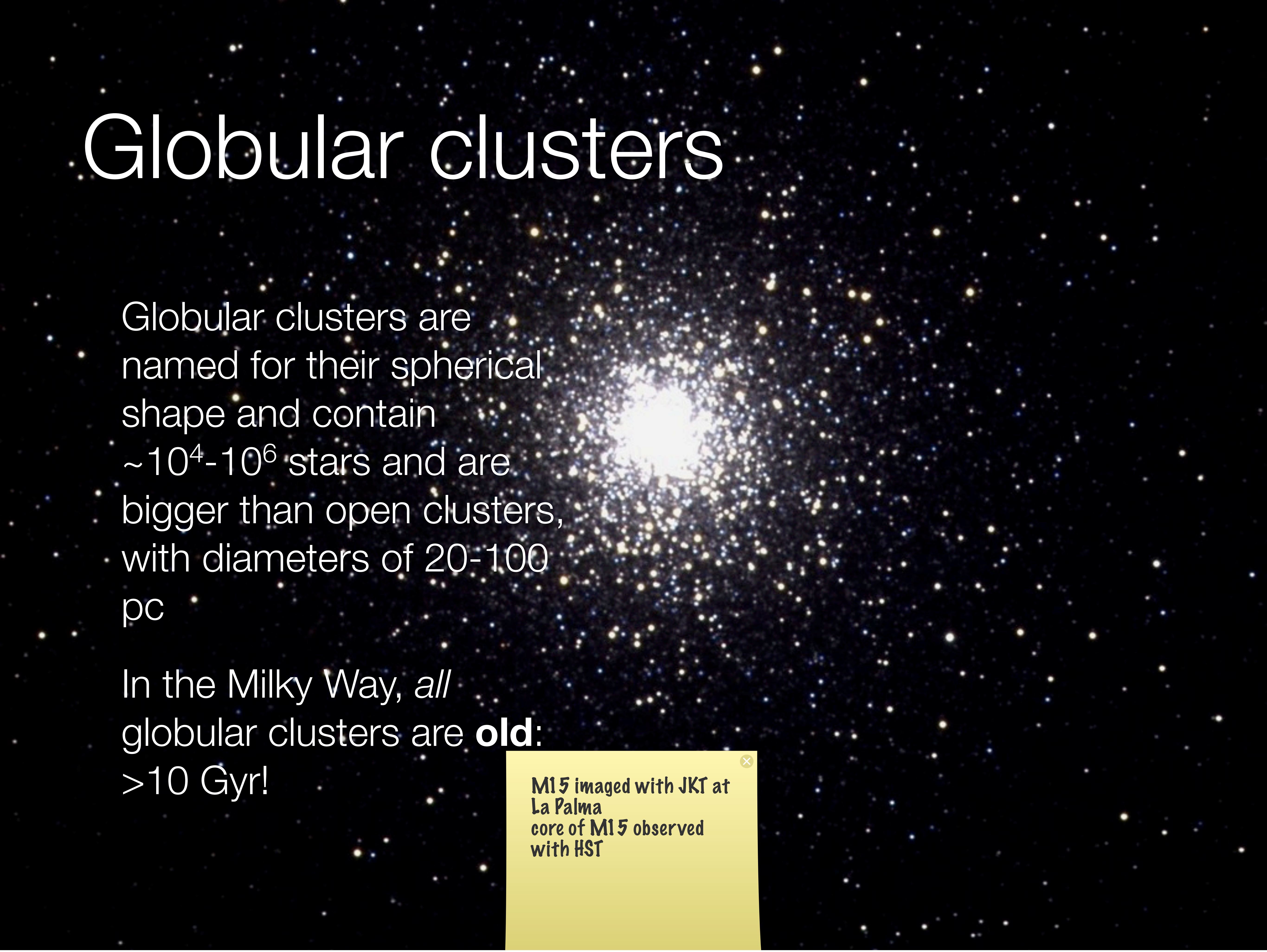

Globular clusters

Globular clusters are named for their spherical shape and contain ~10^4-10^6 stars and are bigger than open clusters, with diameters of 20-100 pc

In the Milky Way, *all* globular clusters are **old**: >10 Gyr!

M15 imaged with JKT at La Palma
core of M15 observed with HST

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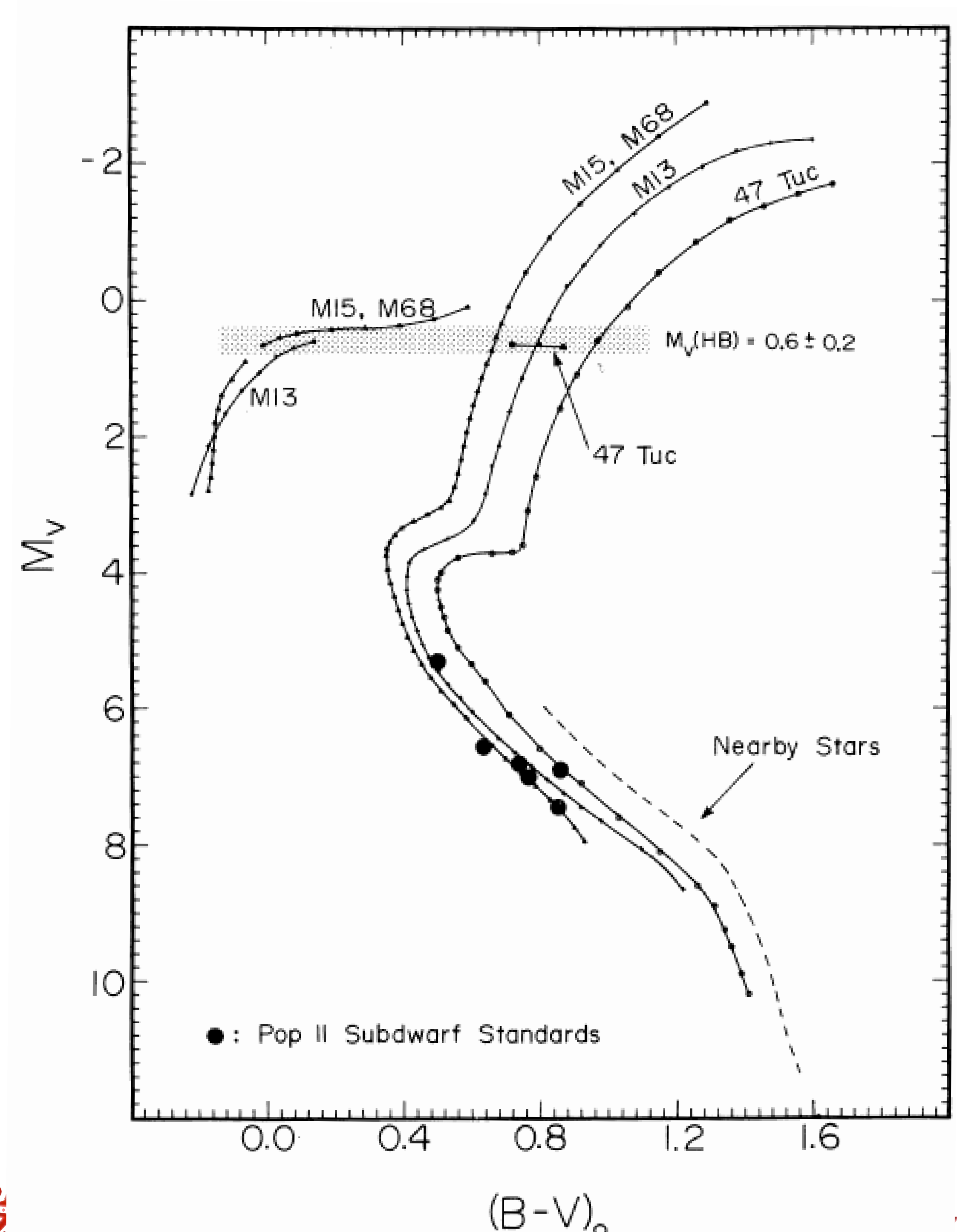
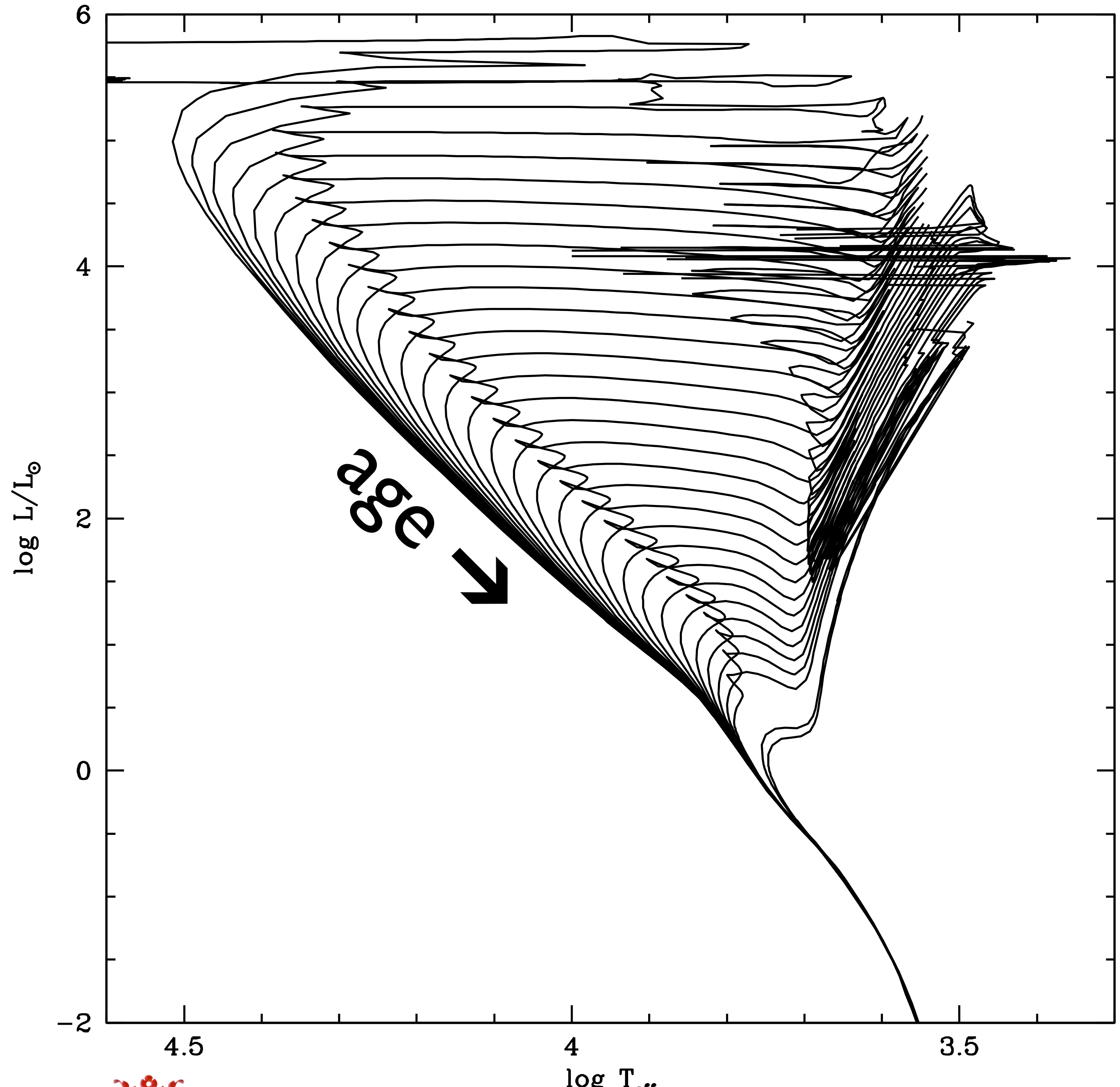
Globular cluster CMDs

- As *all* MW GCs are **old**, their CMDs are very similar. They vary primarily due to *composition differences*
  - “metallicity” [Fe/H]
  - He content
  - variations of other elements
- Age is an important but secondary consideration



Isochrones: "single burst" stellar populations

- Take a set of stellar models all with the **same age** and **same composition** but **different masses**
  - the resulting track in an HR diagram or a CMD is called an *isochrone*



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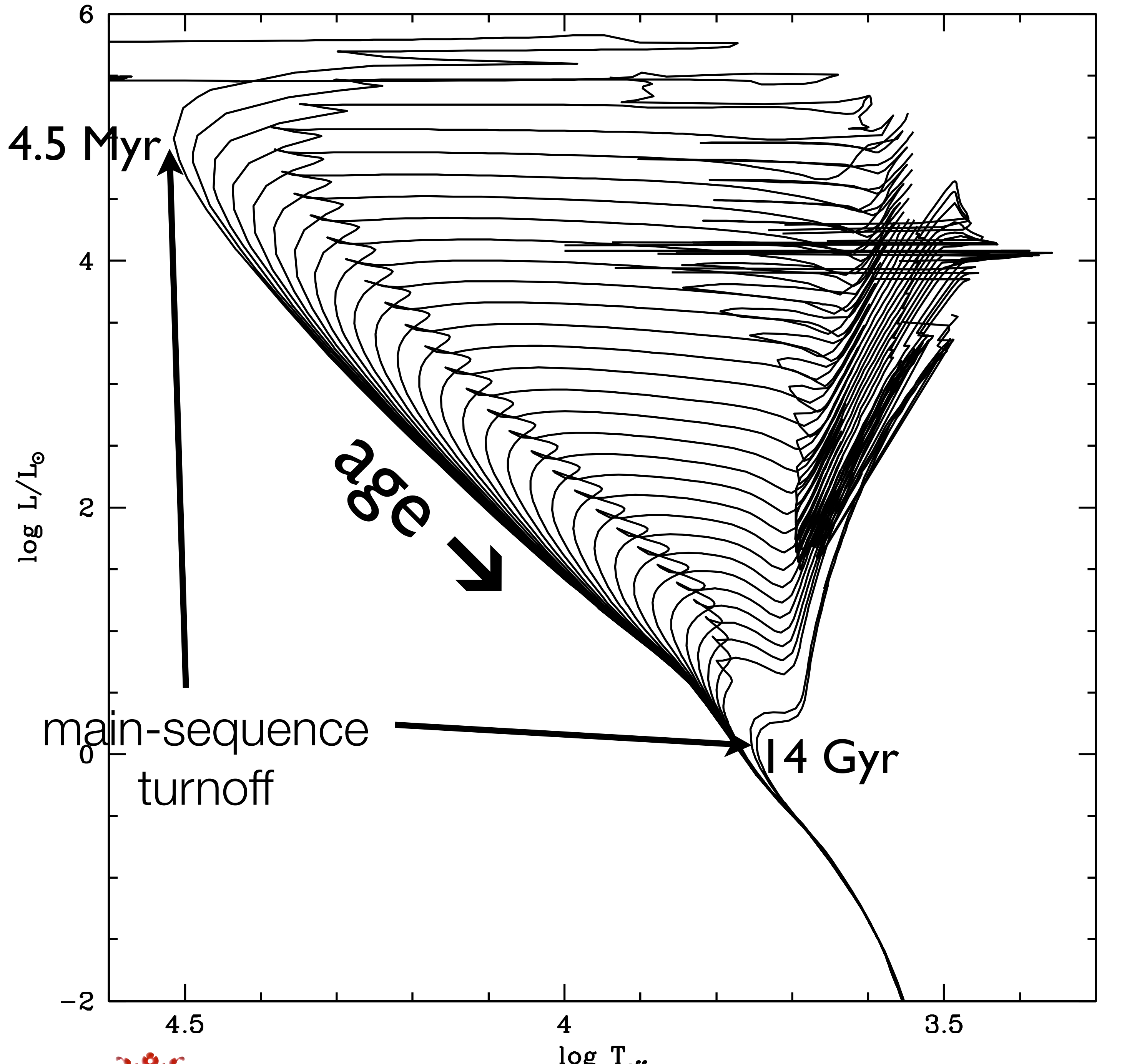
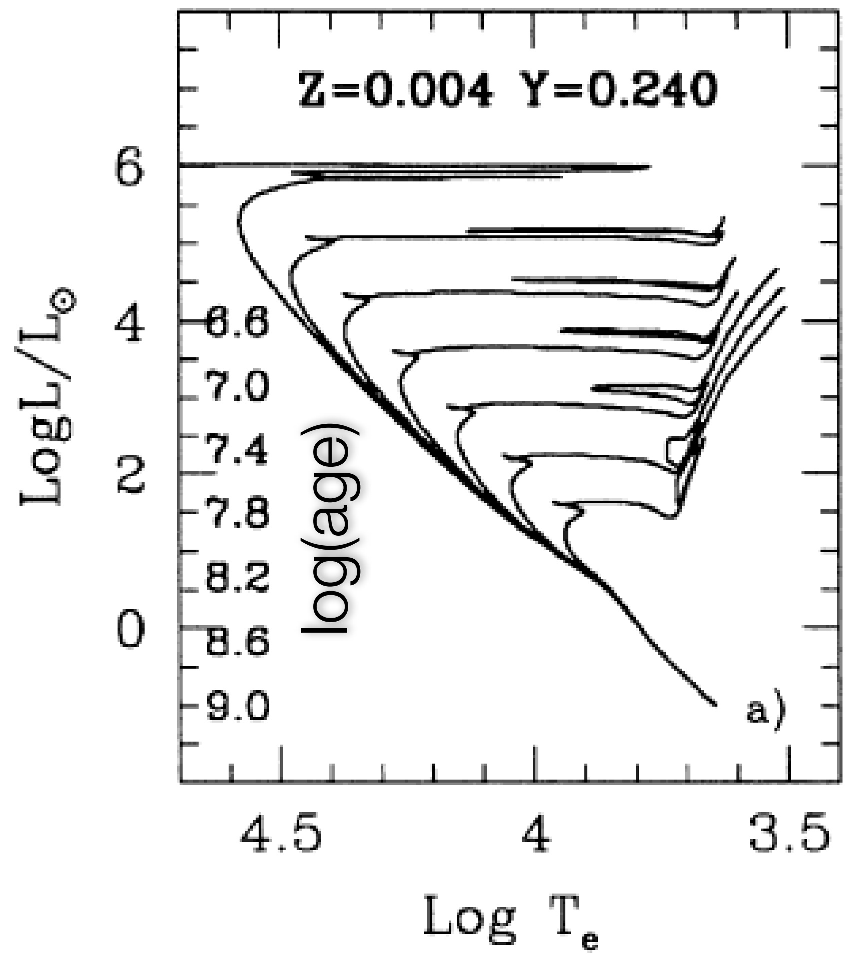
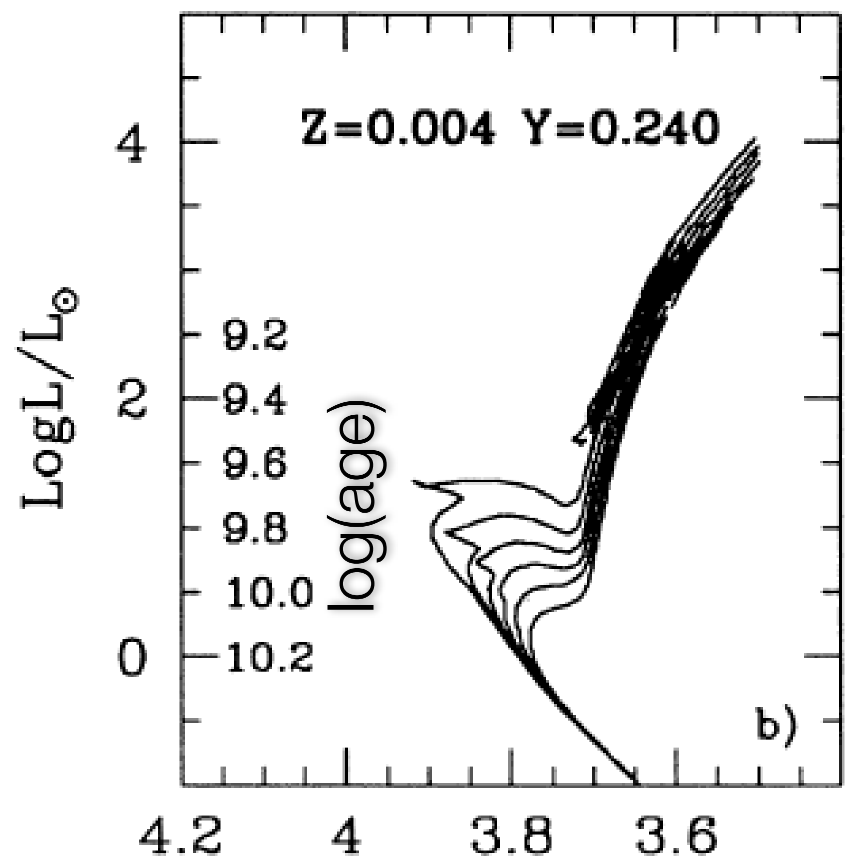


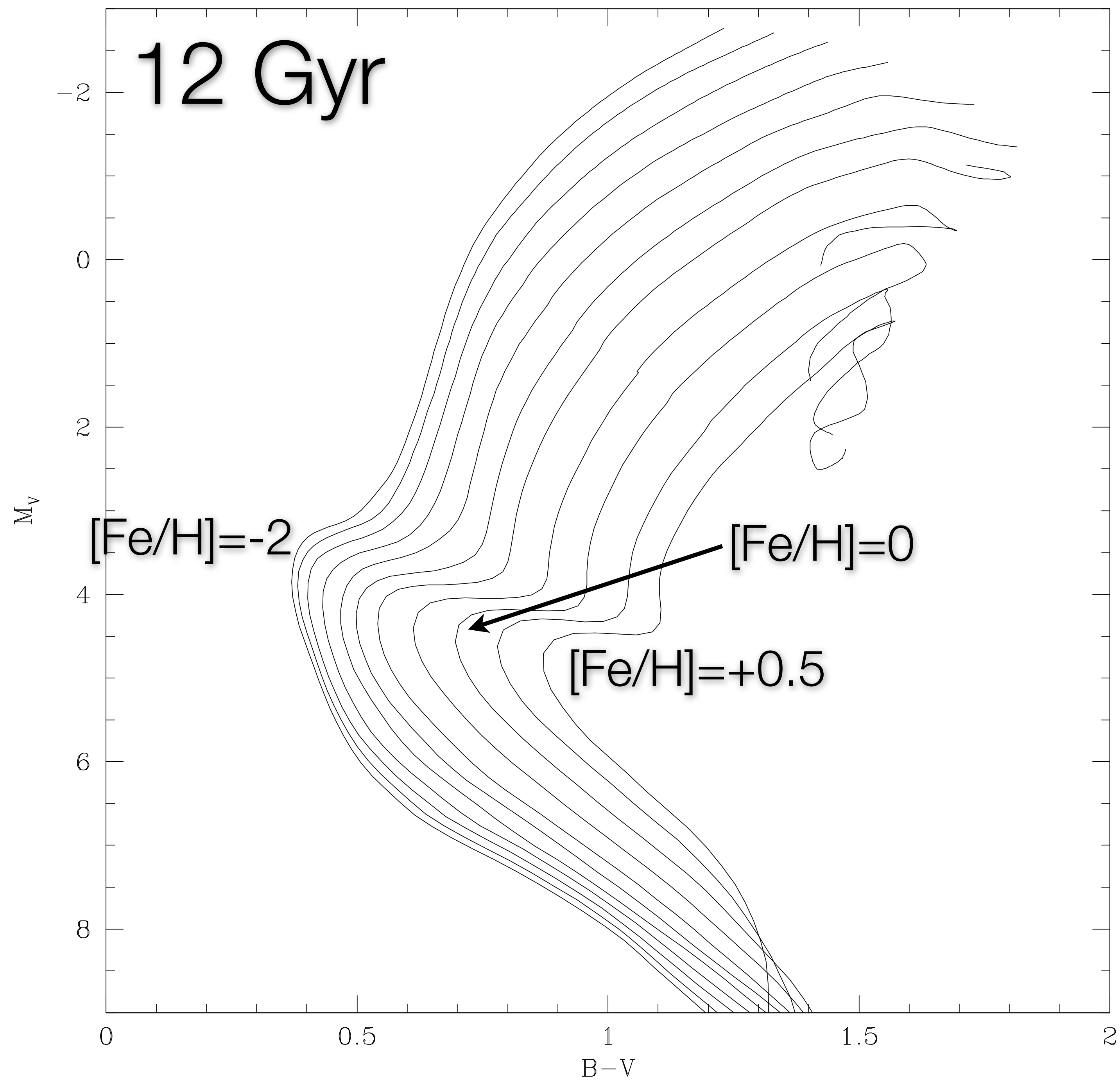
- When populations get old, the isochrones "pile up" at very similar luminosities and temperatures
  - It is easy to determine ages for **young** populations
  - ...but not for **old** populations!
    - "Age dating" GCs is difficult!





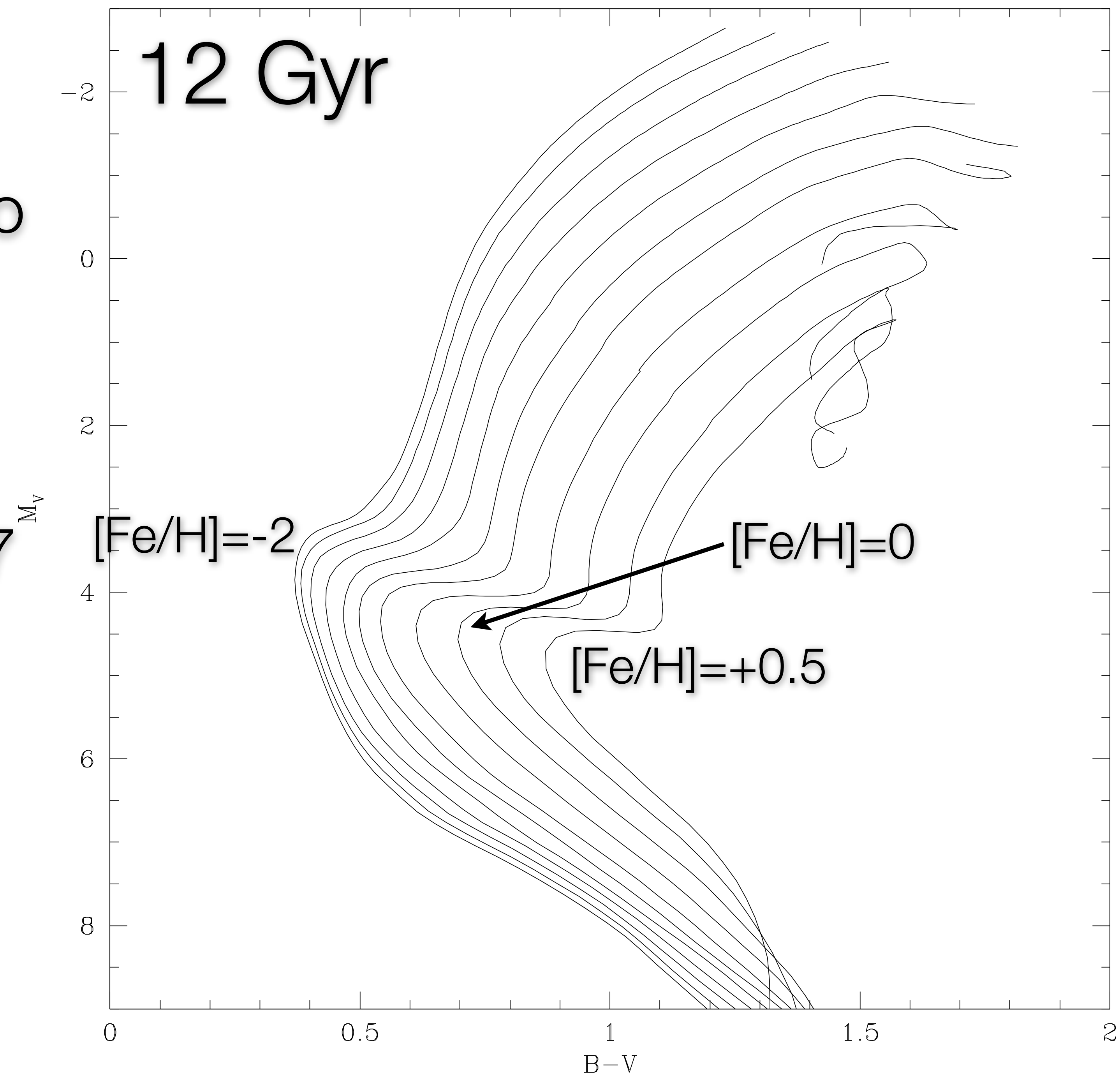
The effect of metallicity

- Stellar temperatures and luminosities also depend on composition at fixed age
  - This is an *opacity* effect: more metals mean more absorption, especially in the blue (it's reradiated into the IR), so stars become cooler (redder) and dimmer



The effect of metallicity

- At fixed L, **Δ**[Fe/H]=–1.7 shifts the unevolved MS (“zero-age MS” or ZAMS) to the blue by **Δ**T_{eff}=+0.06
    - **Δ**(*V*–*K*)=–0.32 mag
- At fixed color, **Δ**[Fe/H]=–1.7 shifts the ZAMS fainter by ~1 magnitude
    - low-metallicity ZAMS stars are thus called “subdwarfs”, because they’re fainter at the same mass



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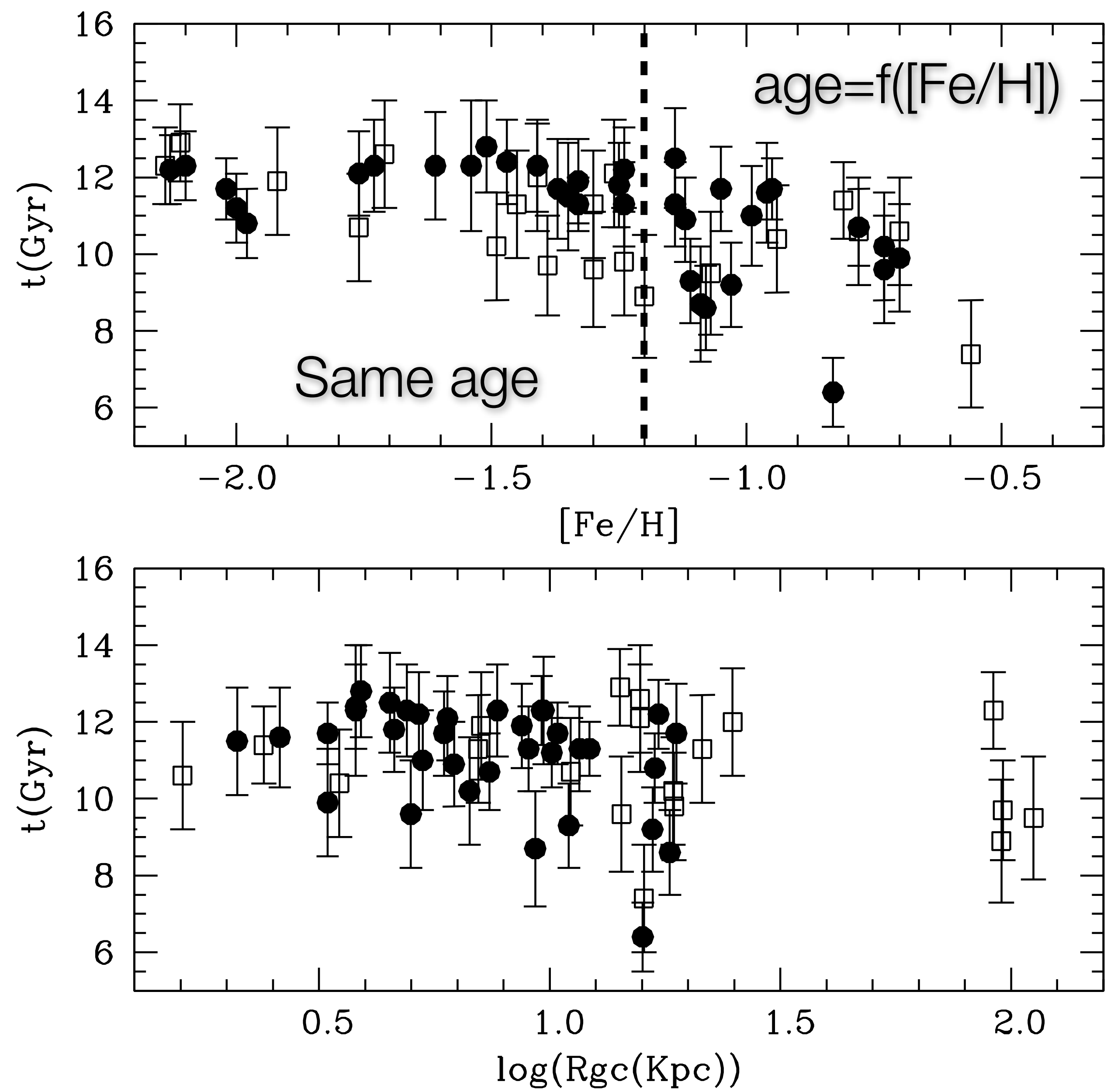
- Putting all of this together, the absolute V magnitude of the MSTO varies as

$$M_V(\mathrm{MSTO}) = 2.7\log(t/\mathrm{Gyr}) + 0.3[\mathrm{Fe/H}] + 1.4$$

- Uncertainties:
    - Difficult to measure the location of the MSTO, as the isochrones are ~vertical
    - Distance errors cause big age errors:
        - 10% distance error ➔ 0.2 mag error in distance modulus ➔ 0.2 mag uncertainty in MSTO magnitude ➔ 20% error in age

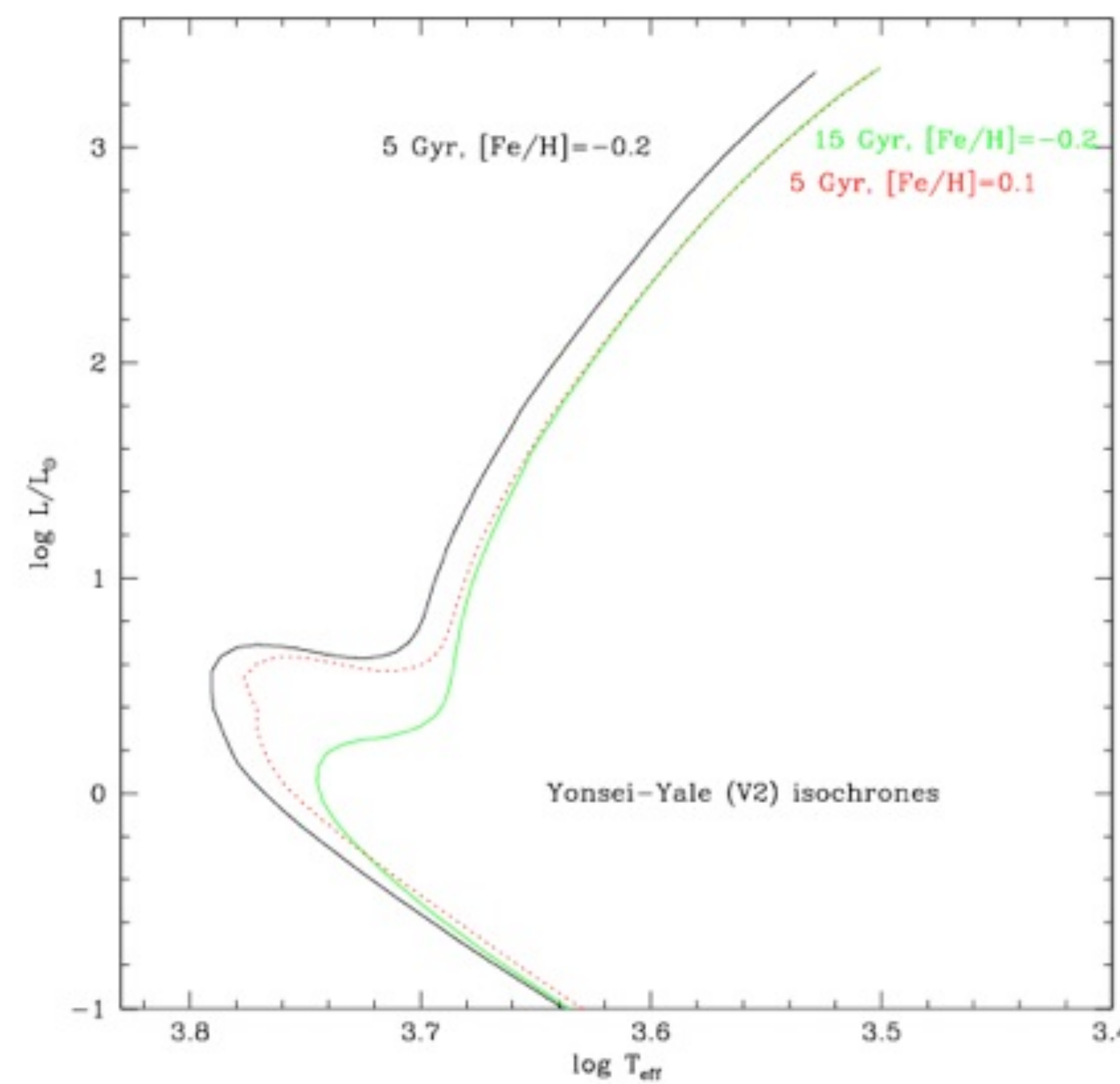
- Globular clusters in the Milky Way are generally very old (>10 Gyr), with only a few younger, metal-rich globular clusters



The age–metallicity degeneracy

- Because **both** age and metallicity affect isochrones, it can be *very* difficult to separate their effects on single-burst stellar populations

- Because the *colors* of an *unresolved* stellar population come from both the giant branch and the MS, colors are **not useful** to determine the ages of "old" populations



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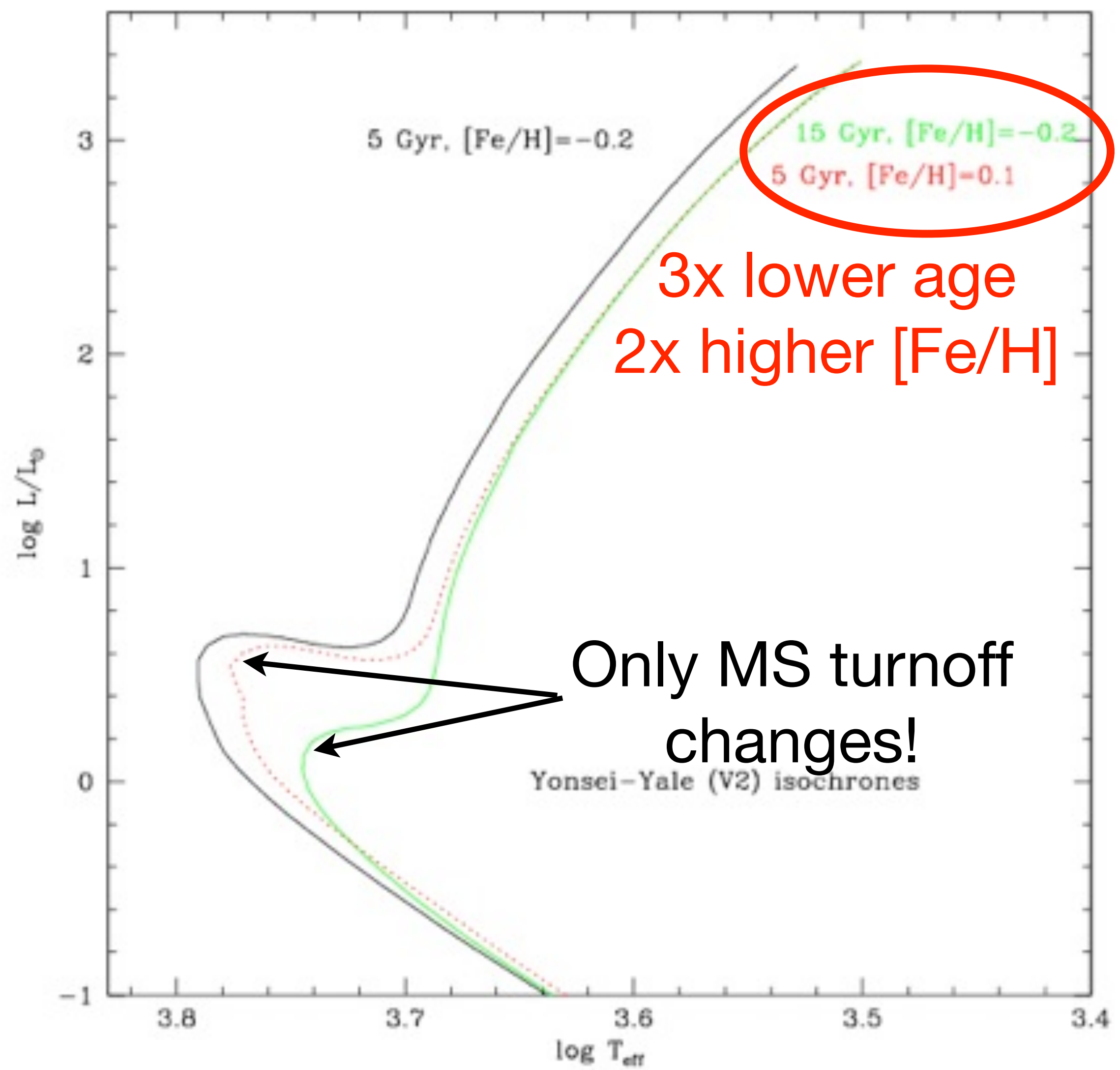

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The luminosity function

- At a given point in the galaxy, the number of stars will vary, with both luminous and faint stars
  - Consider a number dN of stars with absolute magnitudes in the range (M, $M+dM$) in the volume element $d^3\mathbf{x}$ around the point $\mathbf{x}$:

$$dN = \Phi(M, \mathbf{x})\, dM\, d^3\mathbf{x}$$

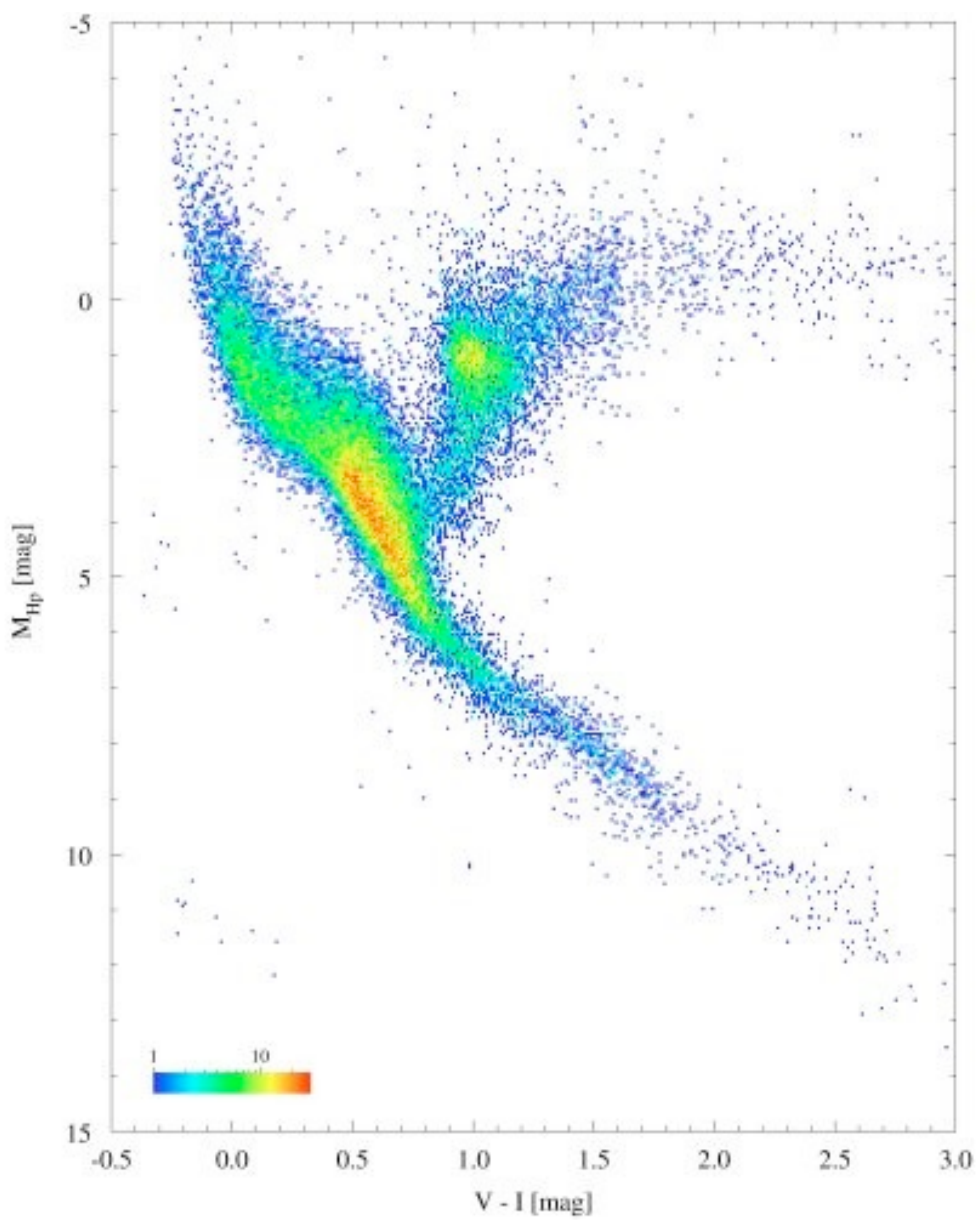
- Now let's assume that the *mix* of stars of different luminosities is independent of location (not always true...):
  - Then we can separate $\Phi(M, \mathbf{x})$ into two functions, $\Phi(M)$ and $\nu(\mathbf{x})$: $dN = [\Phi(M)\, dM][\nu(\mathbf{x})\, d^3\mathbf{x}]$
  - We call $\Phi(M)$ the **luminosity function**, the relative fraction of stars of different luminosities, while $\nu(\mathbf{x})$ is the total number density of stars at **x**.

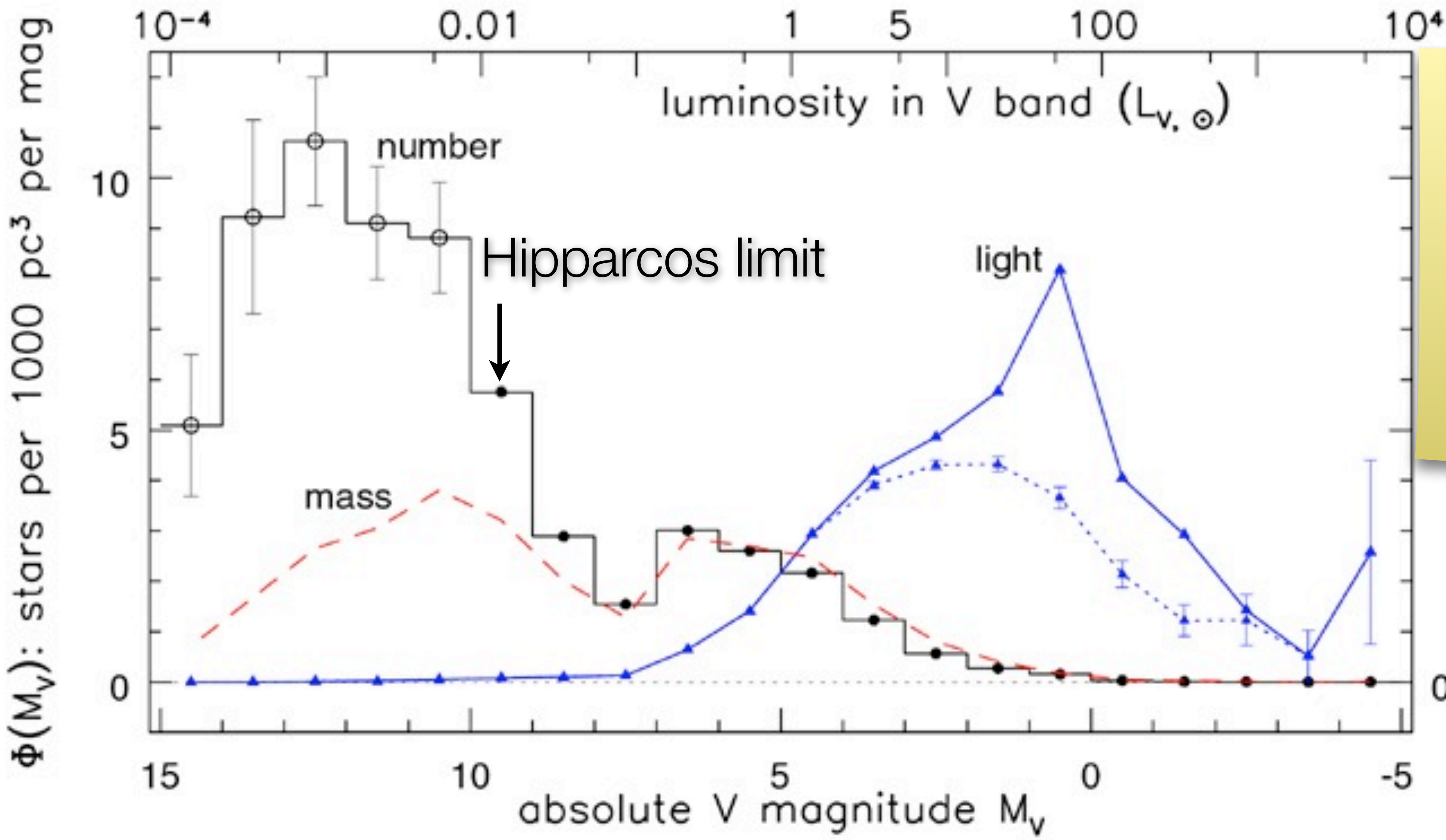
- Note that we will encounter and use the concept of luminosity functions even when we talk about galaxies
  - In this case we generalize $\Phi(M)$ to mean the relative fraction of galaxies with different luminosities and $\nu(\mathbf{x})$ to be the number of density of galaxies

- Let's consider again the local stars observed by Hipparcos within 100 pc
  - Hipparcos was *targeted* on stars with m_V<8
  - So we can construct an LF by writing

$$\Phi(M_V)dM = \frac{dN}{V_{\rm max}}$$

  - where $V_{\rm max}$ is the volume over which stars with M_V could be seen





dots: light from MS stars only
solid: light from all stars

- This is the result (in black). Note the following:
  - Most of the stars are **faint**: peak is at M_V~14
  - If we weight by the *luminosity* of the stars, nearly all of the light is in **bright** stars: peak is at M_V~1
  - If we weight by the *mass* of the stars, most of the stellar mass is in **low-mass** stars: over range 3<M_V<15

- Another interesting result is the **mass-to-light ratio** of stars in the Solar Neighborhood:
  - the *V*-band luminosity density is $0.053 L_\odot \, \mathrm{pc}^{-3}$
  - the mass density is (including white dwarfs) $0.039 M_\odot \, \mathrm{pc}^{-3}$
  - combining, the mass-to-light ratio in solar units is

$$M/L_V \sim 0.67 \, M_\odot / L_\odot$$

  - this is a *lower limit* to the total mass per unit luminosity (in a given band), because we haven't included the dark matter – which we must do in a different way...

- Note that what we've defined here is the **present-day luminosity function** (PDLF)
  - this is the LF we see after the high-mass stars have evolved away

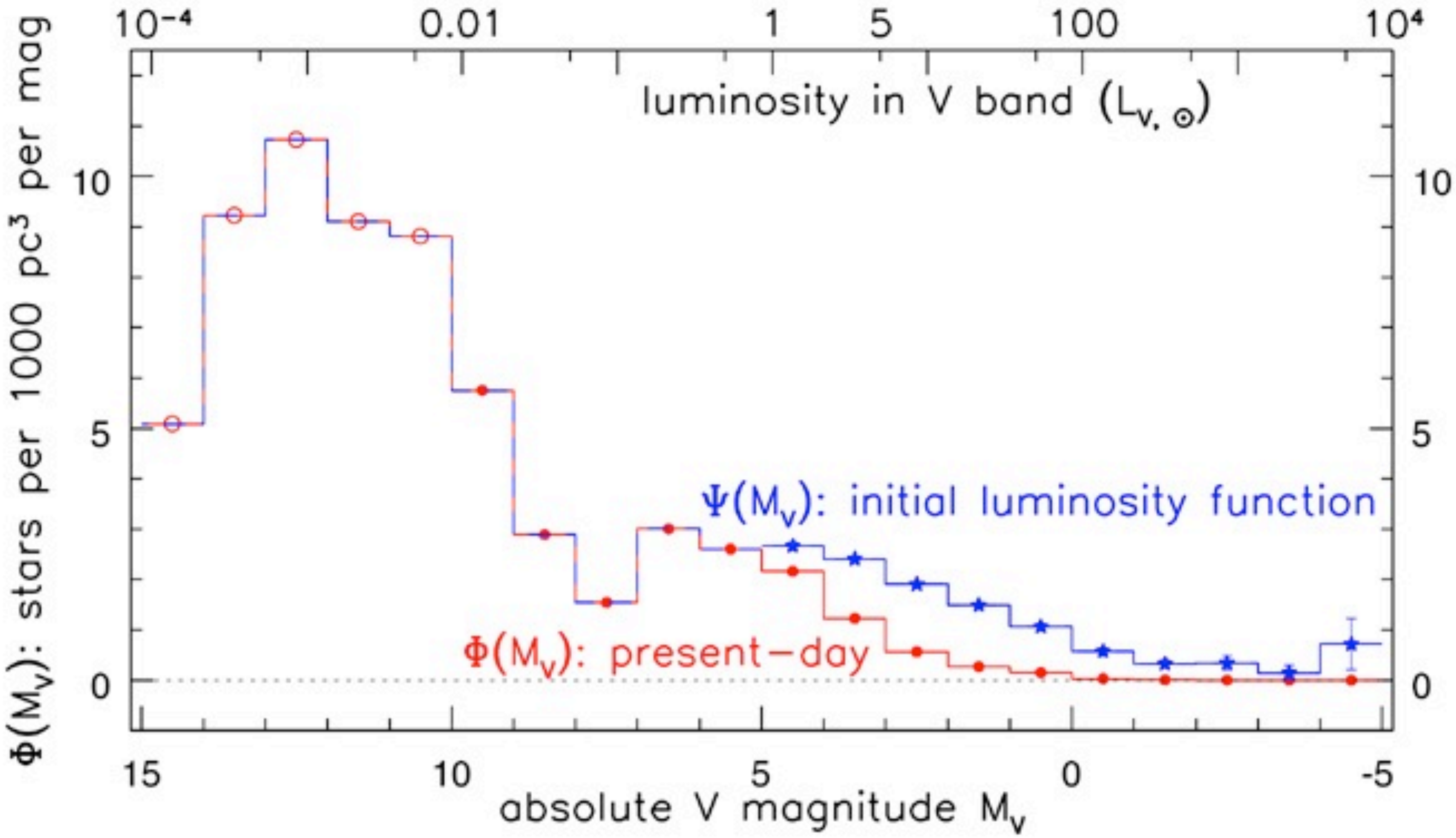
The initial mass function: IMF

- A critical input for stellar population models crucial to understand the formation and evolution of galaxies is the **initial mass function**
  - specifies the distribution of mass in stars *immediately after a star formation event*: the number of stars *dN* with masses between $\mathcal{M}$ and $\mathcal{M}+d\mathcal{M}$ is

$$dN = N_0 \xi(\mathcal{M}) d\mathcal{M}$$

  - We normalize $\xi(\mathcal{M})$ such that N_0 is the total number of solar masses formed in the event:

$$\int \mathcal{M} \xi(\mathcal{M}) d\mathcal{M} = M_\odot$$



- To determine $\xi(\mathcal{M})$, we need to determine *dN* immediately after a star-formation event to get the *initial luminosity function* Φ_0: remember that $dN = \Phi(M)\,dM$
  - if all the stars *just* formed in one burst: no correction

- If that's not the case, but the system has been forming stars at a *constant* rate (roughly true for the MW disk), then

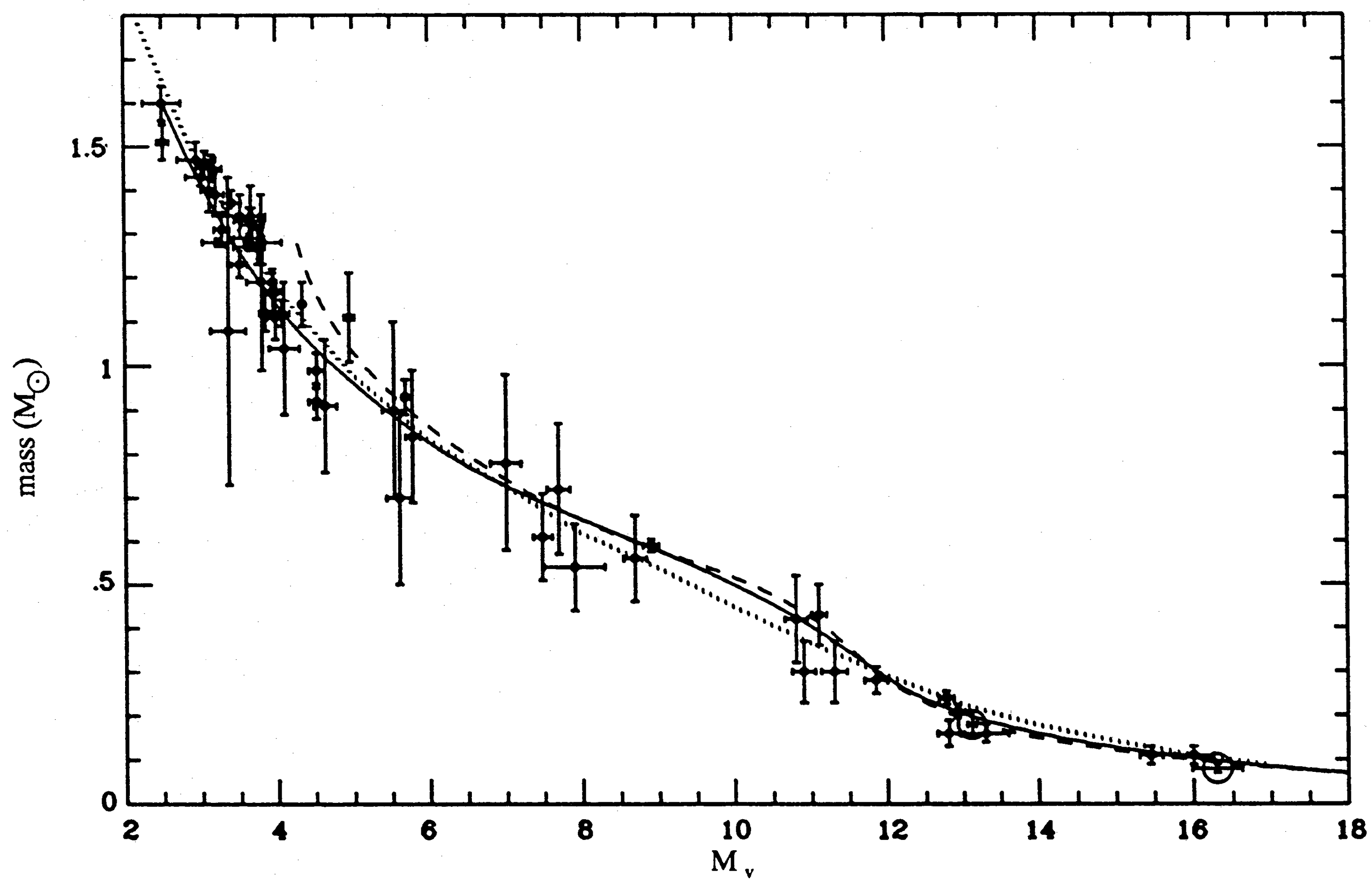
$$\Phi_0(M) = \Phi(M) \times \begin{cases} t/\tau_{\rm MS} & \text{for } \tau_{\rm MS}(M) < t \\ 1 & \text{otherwise} \end{cases}$$

  - The factor $t/\tau_{\rm MS}$ corrects for the fact that we only see stars of magnitude *M* that formed in the last fraction $\tau_{\rm MS}/t$ of the population's life

- Now we can determine ξ($\mathcal{M}$):

$$\xi(\mathcal{M}) = \frac{dM}{d\mathcal{M}} \Phi_0[M(\mathcal{M})]$$

- The function $M(\mathcal{M})$ specifies the relation between stellar mass and absolute magnitude
    - unfortunately, you also need the derivative of this function...



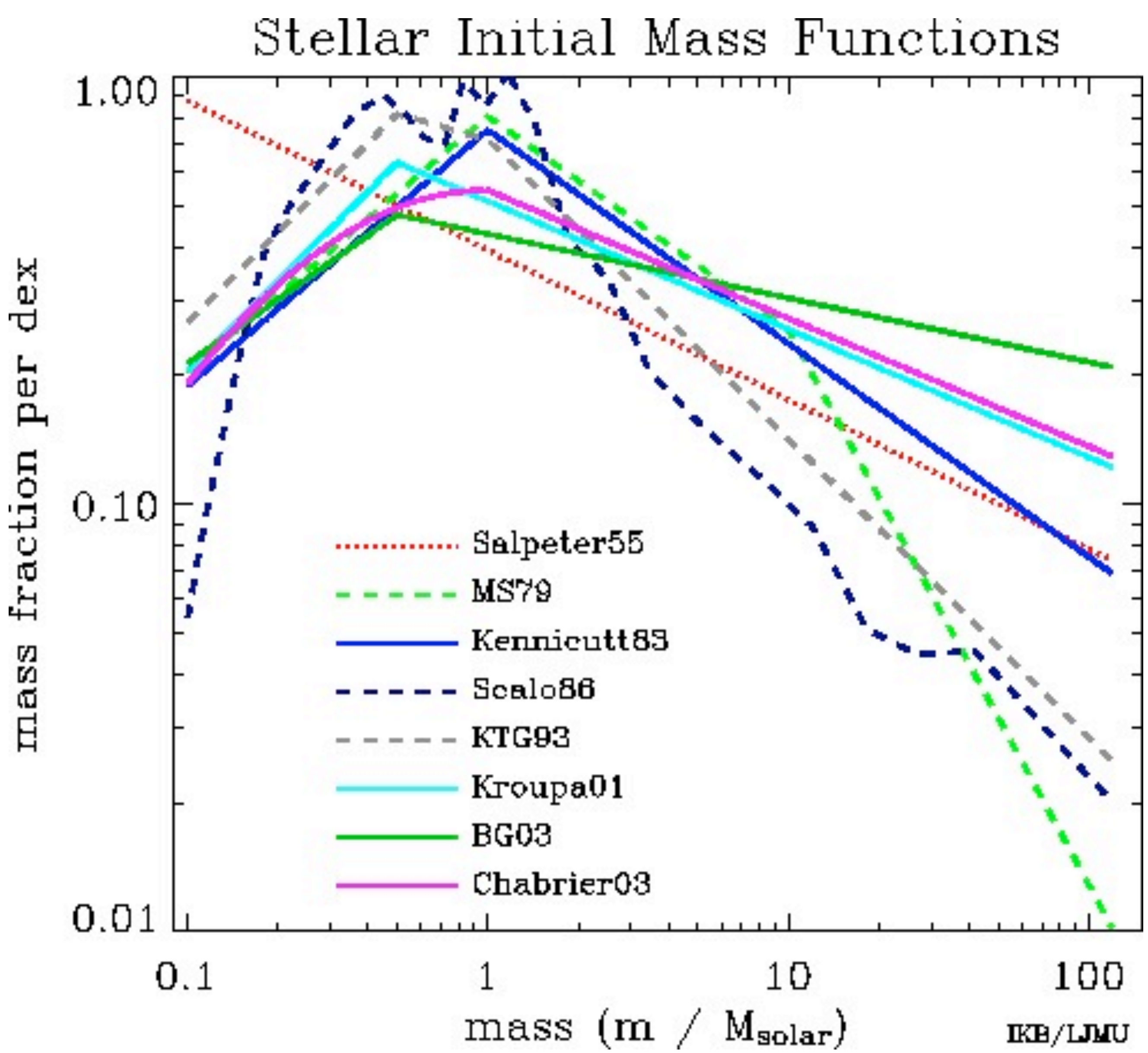
- The simplest (and still commonly used) IMF is that inferred by Salpeter (1955), which is a *power-law*:

$$\xi(\mathcal{M}) \propto \mathcal{M}^{-2.35}$$

  - note that you often see the exponent as –1.35, but this is a different definition of the IMF, in log $\mathcal{M}$ instead of (linear) $\mathcal{M}$

- In reality, we know that the low-mass IMF must be *flat* or even *decline* with decreasing mass, at least in the Solar Neighborhood

- This has led to a number of different parameterizations of the IMF...
  - Salpeter, Kroupa, and Chabrier are the most commonly-used IMFs today



- The big question: **is the IMF universal?**
  - Nearly everyone assumes so... but this may not be true....

The evolution of stellar populations

- Stars with masses <1.5 $M_\odot$ live for >2.5 Gyr (at solar abundance) and put out most of their energy **after** the MSTO
  - for instance, a 1 $M_\odot$ star puts out E_{MS}~10.8 $L_\odot$ Gyr on the MS and E_{GB}~24 $L_\odot$ Gyr in the post-MS (RGB, HB, AGB) phases

- If stars of mass M emit a total energy of E_{GB} on the giant branch, then the luminosity of the population will be

$$L \approx \left(E_{\rm GB} \frac{dN}{d\mathcal{M}} \right)_{\mathcal{M}_{\rm GB}} \left| \frac{d\mathcal{M}_{\rm GB}}{dt} \right|$$

this is the energy released on the giant branch per star times the change in the number of stars per unit time

- For these low-mass stars, the MS lifetimes are

$$\frac{\tau_{\rm MS}}{10\,{\rm Gyr}} \approx \left(\frac{\mathcal{M}}{M_\odot} \right)^{-3}$$

Note that GB lifetimes are ~10% of MS lifetimes, so the range of masses on the GB is small --- roughly that of the MSTO plus a very small amount

- or

$$\frac{\mathcal{M}_{\rm GB}}{M_\odot} \approx \left(\frac{\tau_{\rm MS}}{10\,{\rm Gyr}} \right)^{-1/3}$$

- and so

$$\frac{d\mathcal{M}}{dt} \approx -\frac{1}{3} \left(\frac{\mathcal{M}_{\rm GB}}{M_\odot} \right)^4 \left(\frac{M_\odot}{10\,Gyr} \right)$$

- Now, if we take the IMF to be a power-law with $\alpha \leq -2.35$ around $\mathcal{M}_{\mathrm{GB}}$, then $dN/d\mathcal{M} \approx K(\mathcal{M}/M_\odot)^{-\alpha}$
- And the luminosity is then

$$L \approx \frac{K M_\odot E_{\mathrm{GB}}(\mathcal{M}_{\mathrm{GB}})}{30\,\mathrm{Gyr}} \left(\frac{\mathcal{M}_{\mathrm{GB}}}{M_\odot}\right)^{4-\alpha}$$

- Finally, we can differentiate L to find

$$\frac{d\ln L}{dt} \approx \left[\frac{d\ln E_{\mathrm{GB}}}{d\ln \mathcal{M}_{\mathrm{GB}}} + (4-\alpha)\right] \frac{d\ln \mathcal{M}_{\mathrm{GB}}}{dt}$$

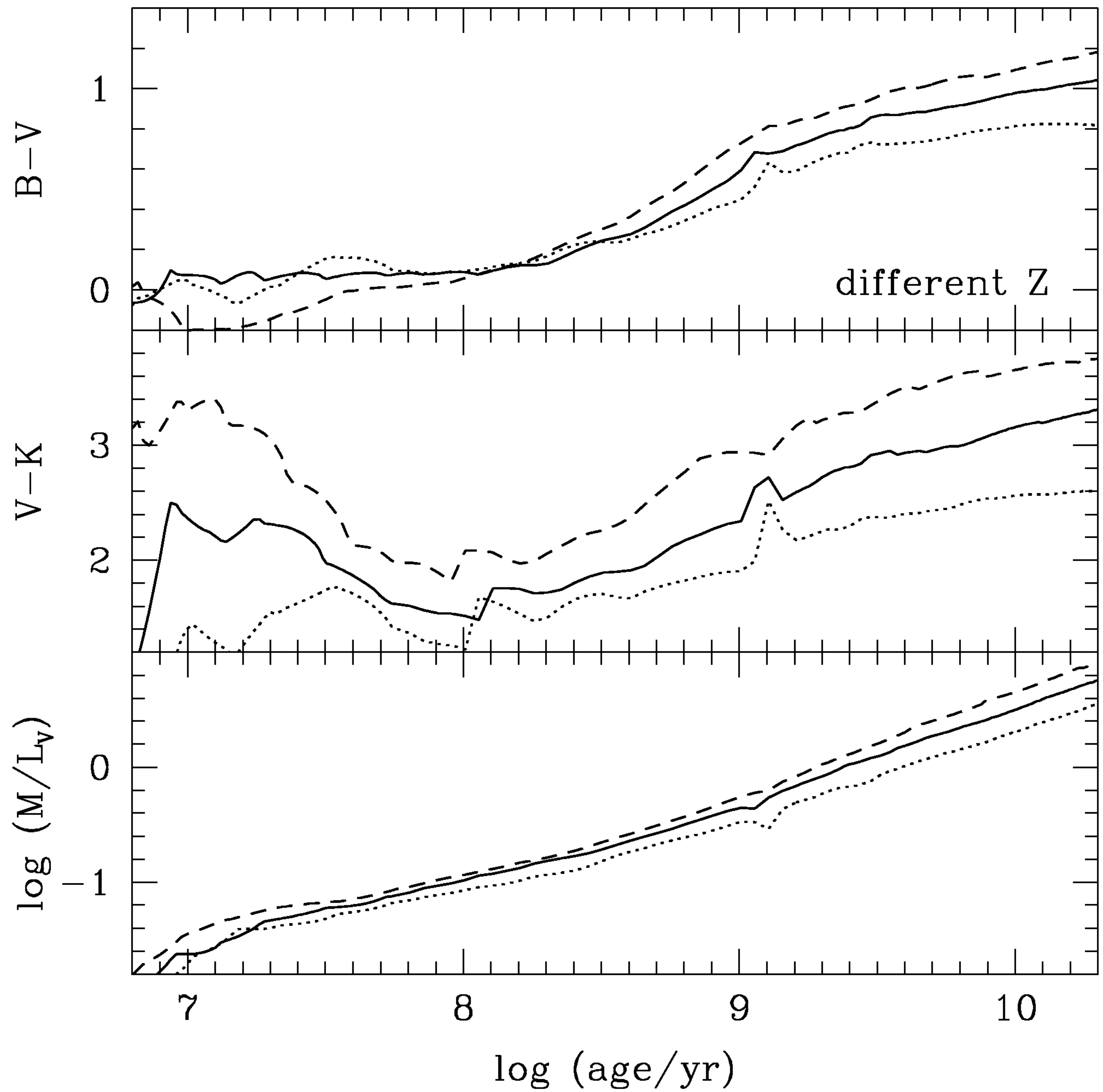
$$\approx 0.3\alpha - \left[1.3 + 0.3\frac{d\ln E_{\mathrm{GB}}}{d\ln \mathcal{M}_{\mathrm{GB}}}\right]$$

take the log of the 2nd eq. on the previous page...

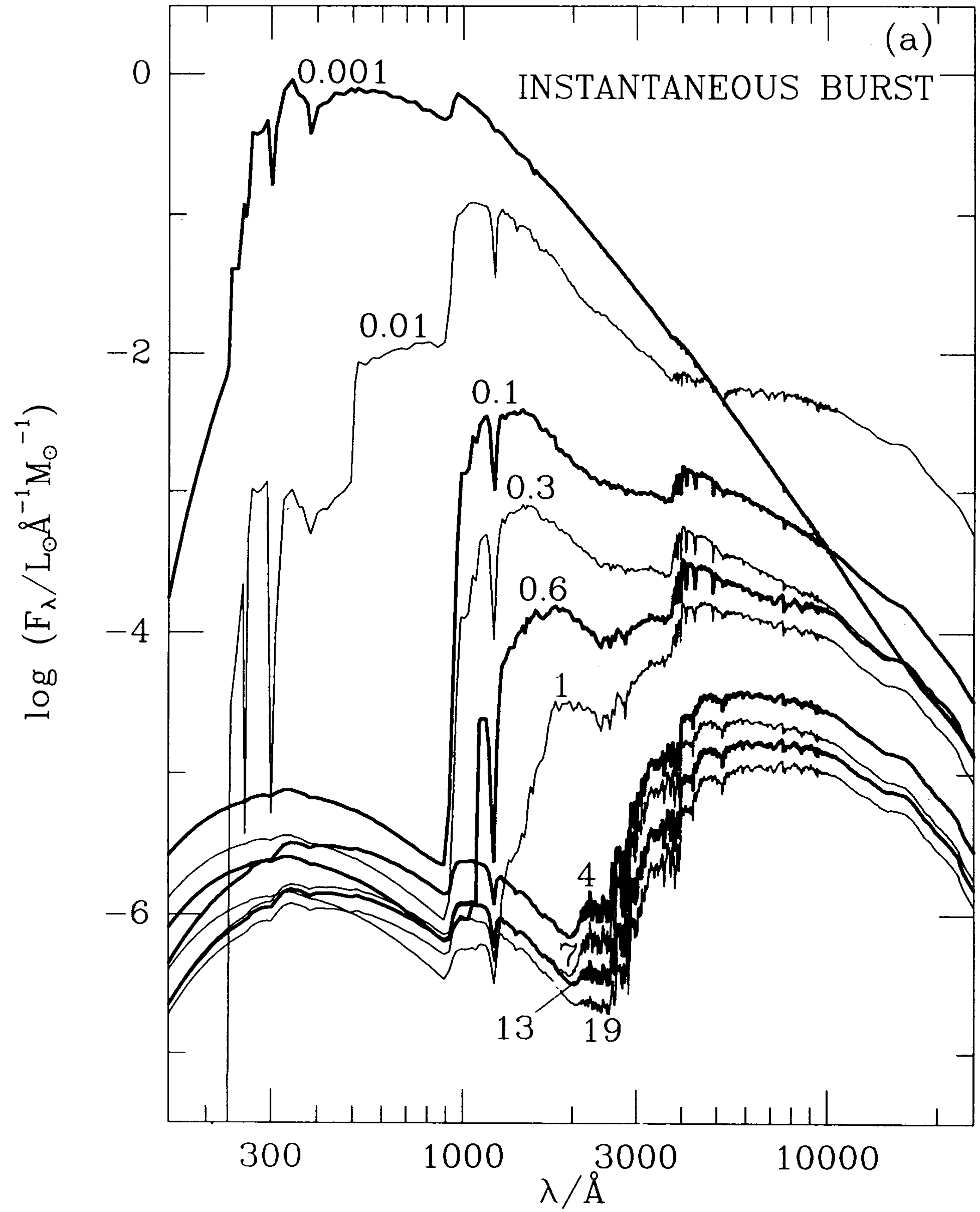
- This is strikingly close to the expression found by Tinsley & Gunn (1976): $(d\ln L/dt) \approx 0.3\alpha - 1.3$

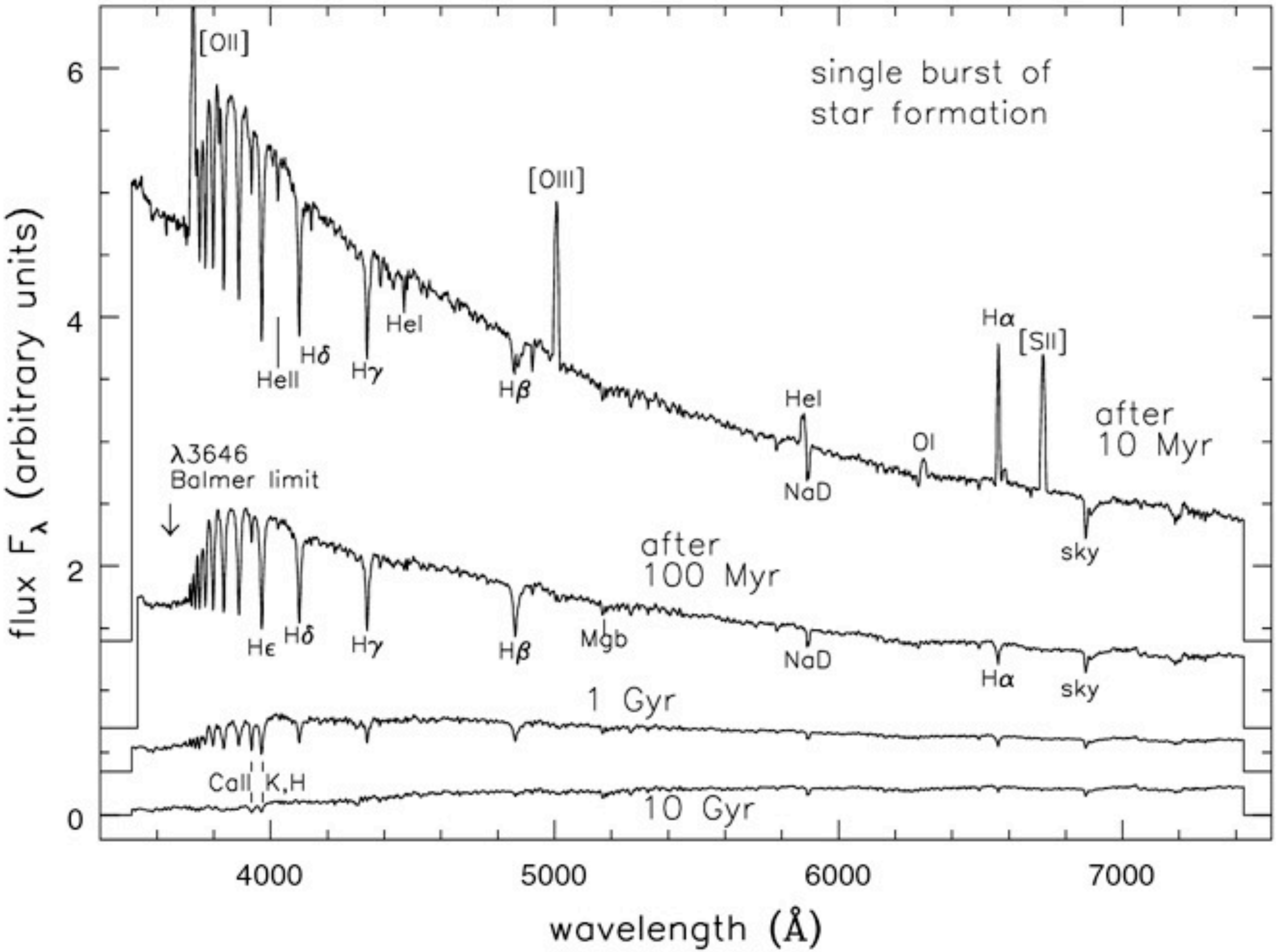
- So as long as $\alpha<4$, then *the luminosity of a coeval population* ***decreases with time***
  - in other words, in the absence of star formation, galaxies (and star clusters) get fainter with time!

- Stellar population models support this (using Salpeter IMFs)
    - note also that populations get redder with age, as expected!



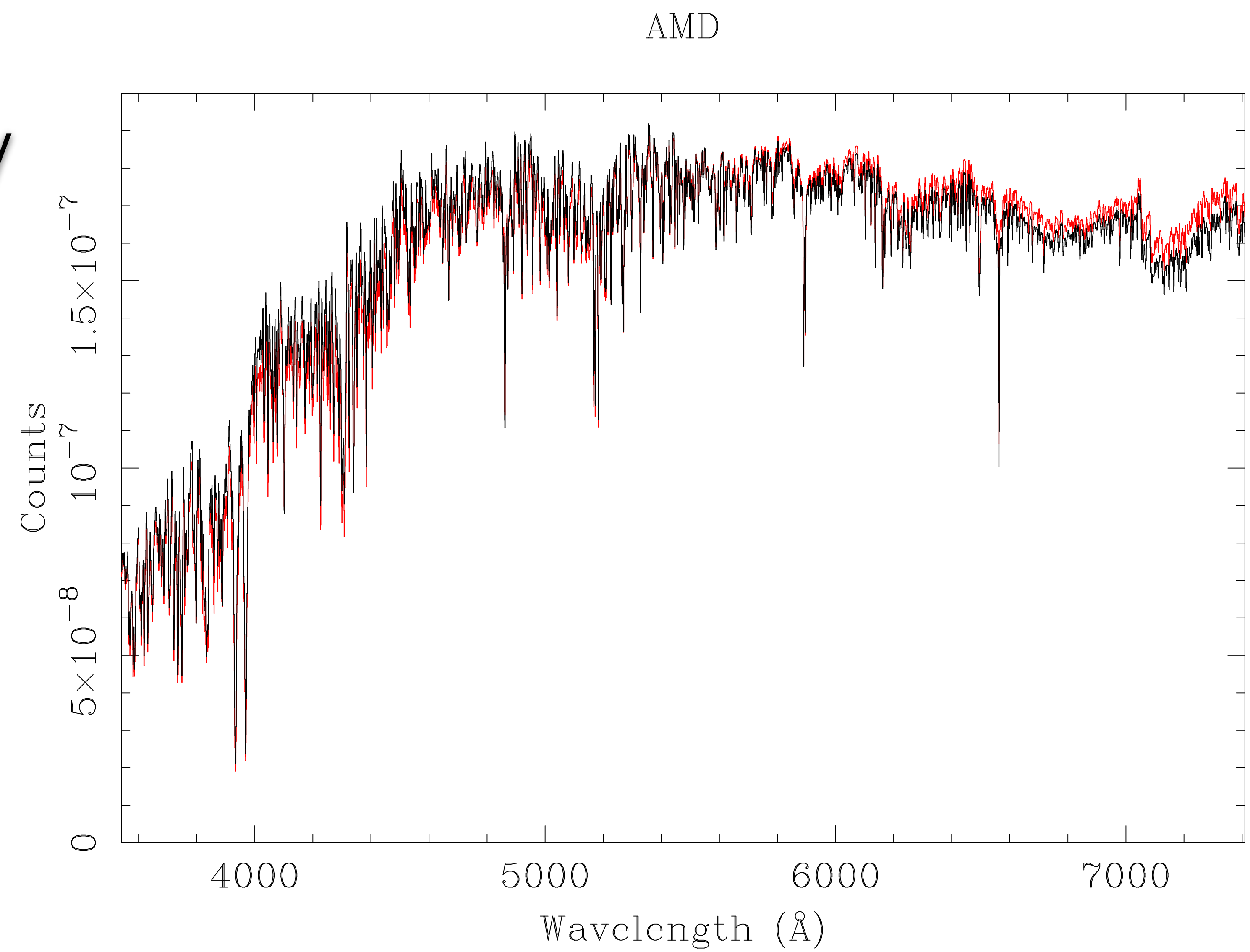
* Spectra do the same thing, of course...





single burst of
star formation
[OII]
[OIII]
HeII
Hδ
Hγ
HeI
Hβ
Hα
[SII]
HeI
OI
NaD
after
10 Myr
sky
λ3646
Balmer limit
after
100 Myr
Hε
Hδ
Hγ
Hβ
Mgb
NaD
Hα
sky
1 Gyr
CaII K,H
10 Gyr
flux Fλ (arbitrary units)
wavelength (Å)
0
2
4
6
4000
5000
6000
7000

- But watch out for the dreaded age–metallicity degeneracy!
- Two spectra with
  - 5 Gyr, [Fe/H]=-0.4
  - 15 Gyr, [Fe/H]=-0.7
- Can you tell the difference?



Vazdekis MILES models

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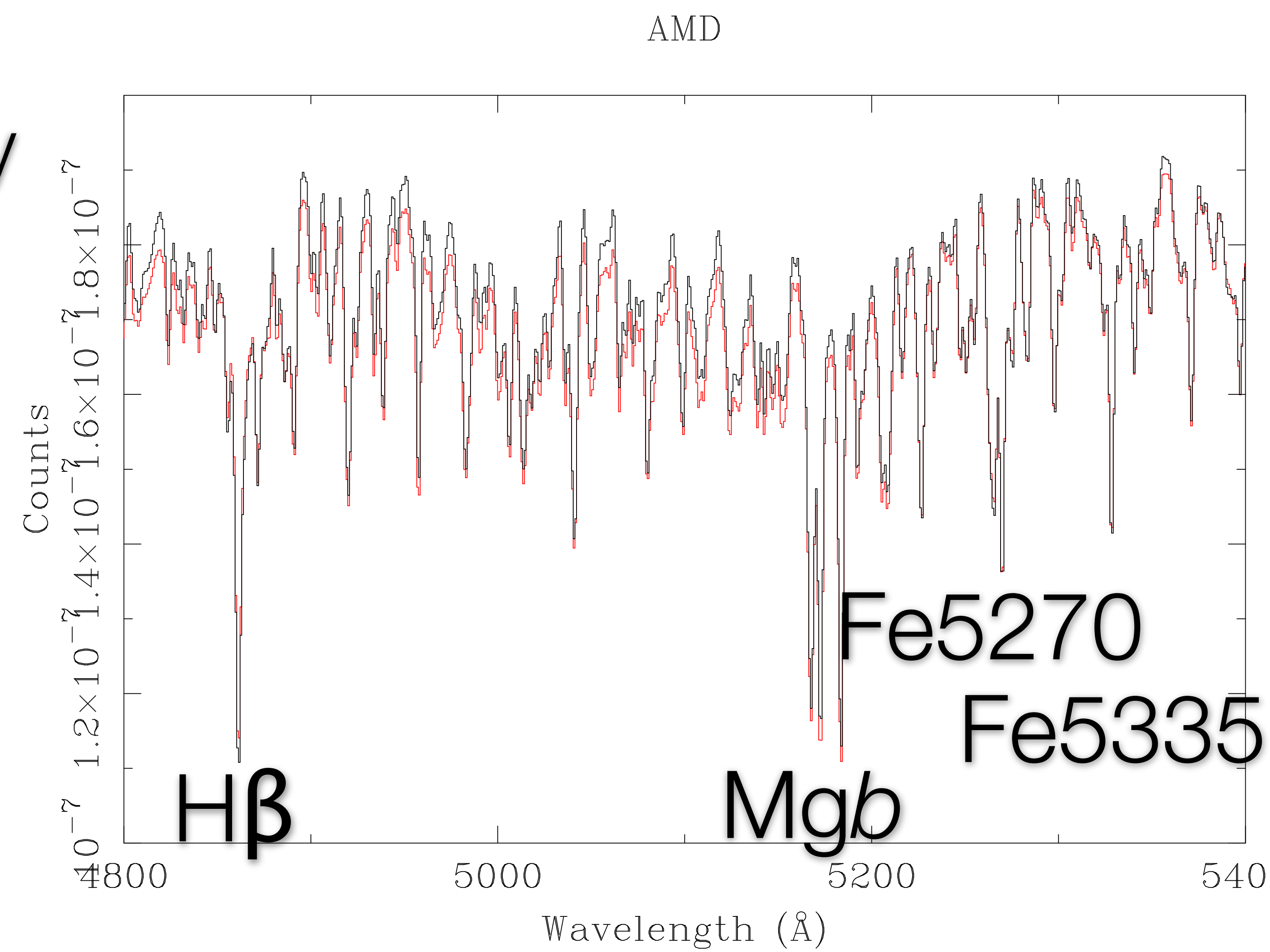

Vazdekis MILES models