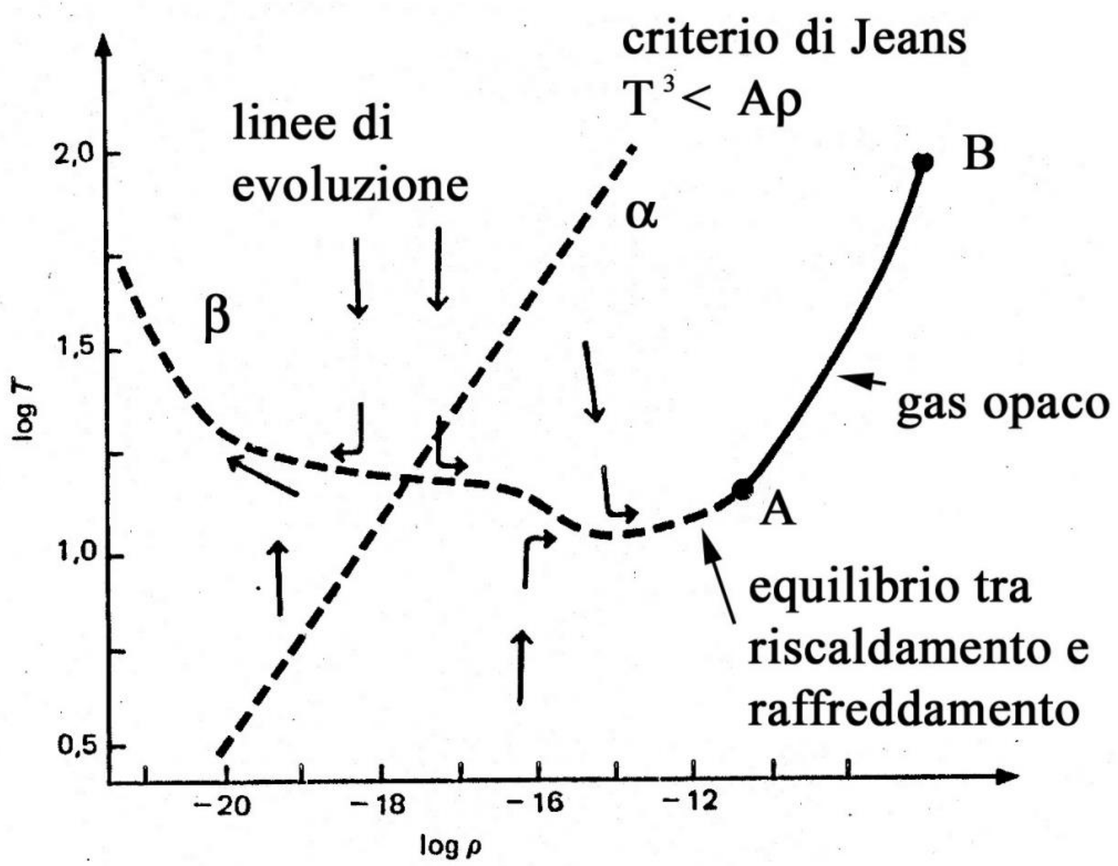
Contraction of molecular clouds

La curva β definisce il bilancio tra processi di raffreddamento e di riscaldamento

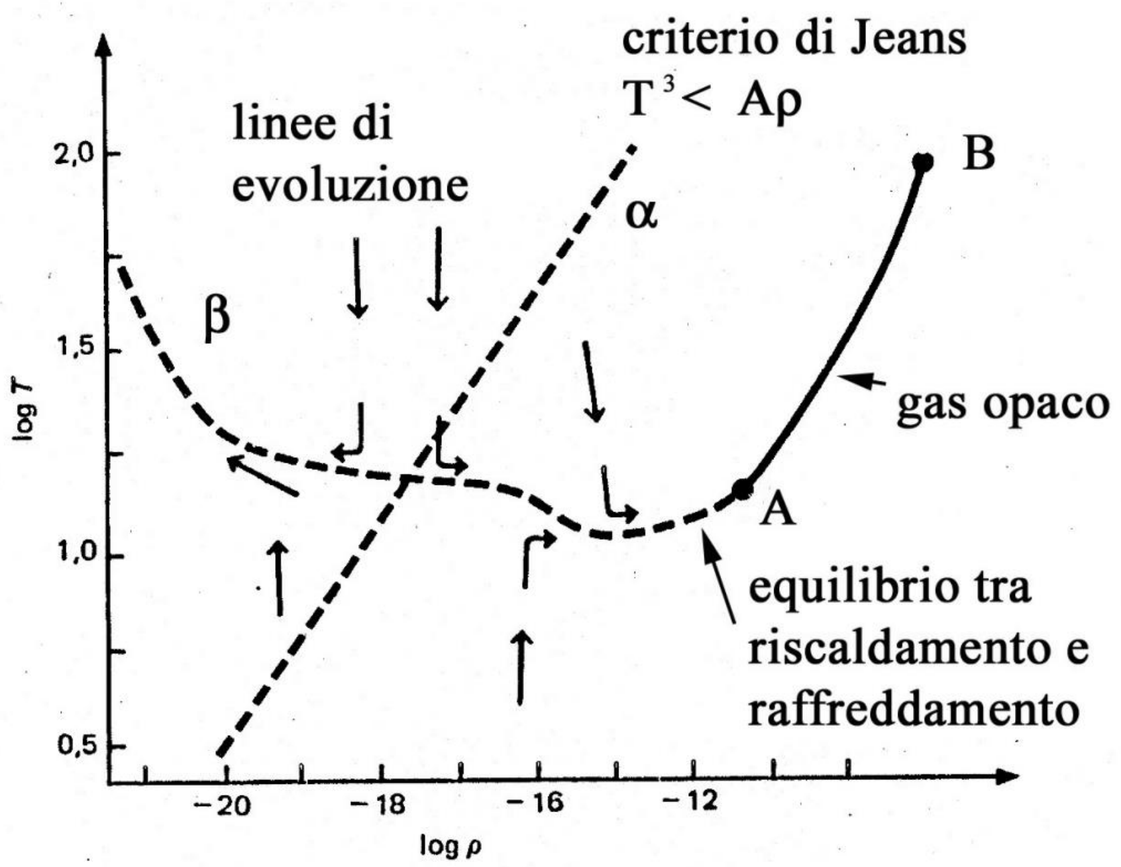
The contracting cloud gives rise to a low-temperature spherical gas mass, formed by the condensation of interstellar matter as a consequence of gravitational instability.
Initially, it is an isothermal spherical mass (the material being transparent at first) with finite dimensions; such a configuration is unstable, and the protostar continues to contract until it reaches point A.



Curves on evolution of clouds in contraction according to Hayashi

Contraction of molecular clouds

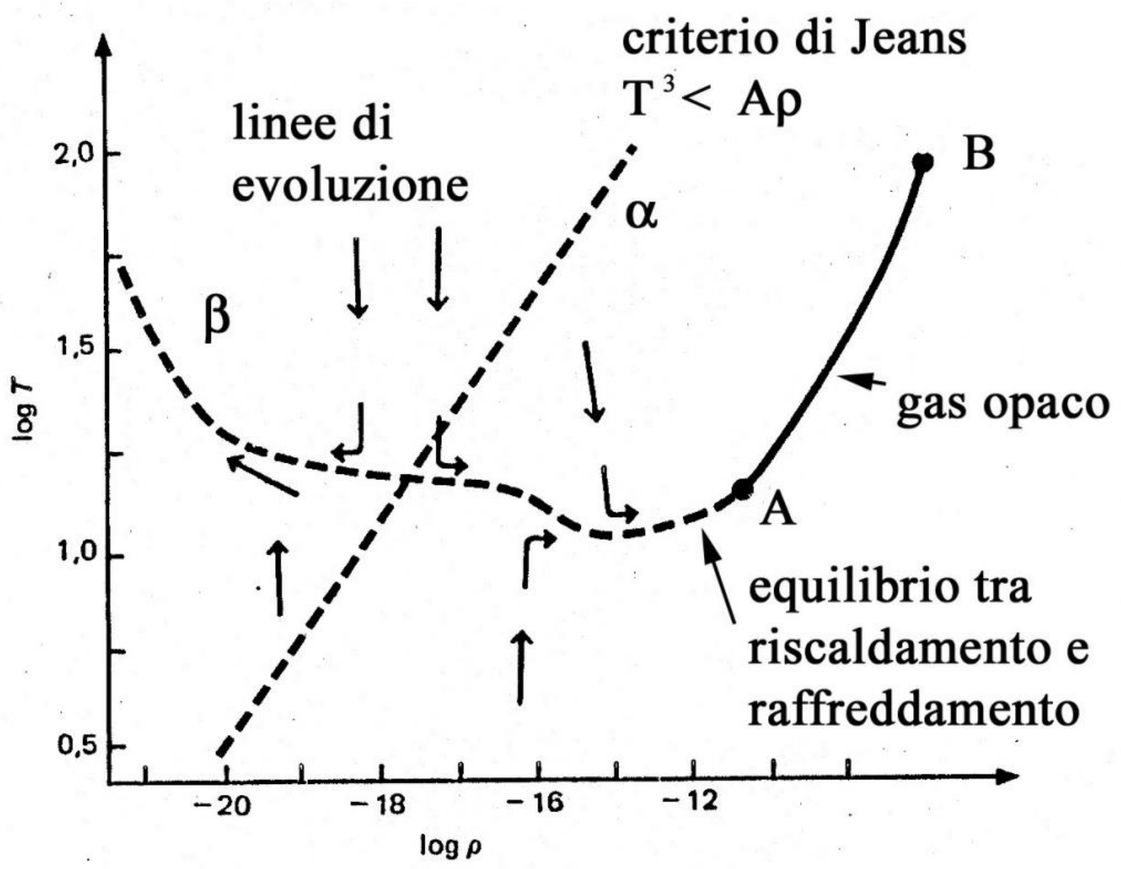
In the case of contraction, the internal pressure is insufficient to counteract the gravitational force, and the system evolves essentially under free-fall conditions at constant temperature. Consequently, the timescale is short: for a cloud of 1 M⊙ with a radius of 10^6 R⊙, the contraction time is approximately 10^4 years.



Curves on evolution of clouds in contraction according to Hayashi

Contraction of molecular clouds

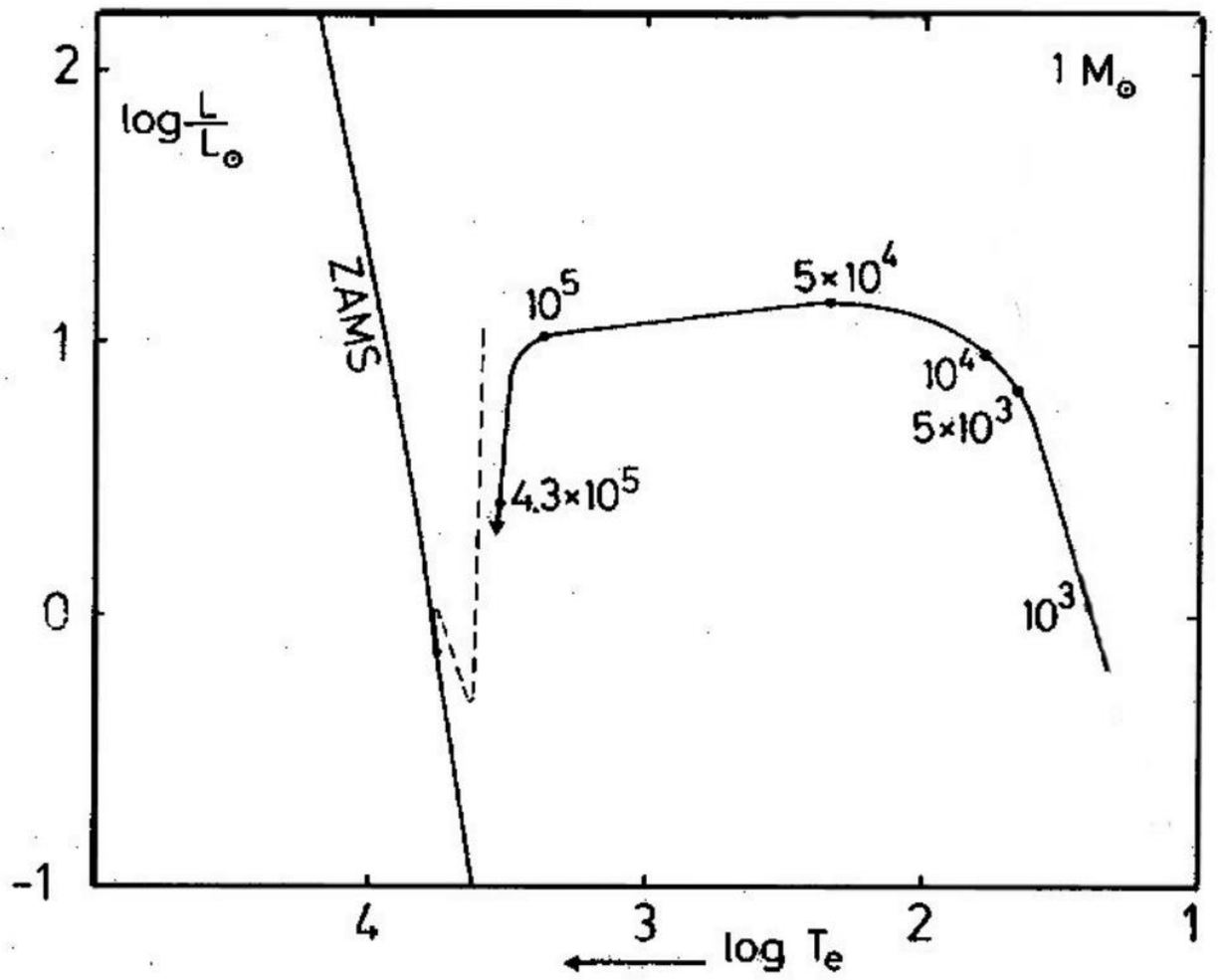
At point A, the cloud becomes opaque to radiation and continues its evolution through a slower phase of adiabatic contraction. As predicted by the virial theorem, it begins to heat up.
The increase in internal pressure occurs mainly in the central regions of the cloud, which reach hydrostatic equilibrium and generate a shock wave in the outer material, still contracting at free-fall velocity.
The propagation of the shock wave ionizes the material and “ignites” the cloud — at this stage, it can be referred to as a **protostar**.



Curves on evolution of clouds in contraction according to Hayashi

Proto stellar evolution

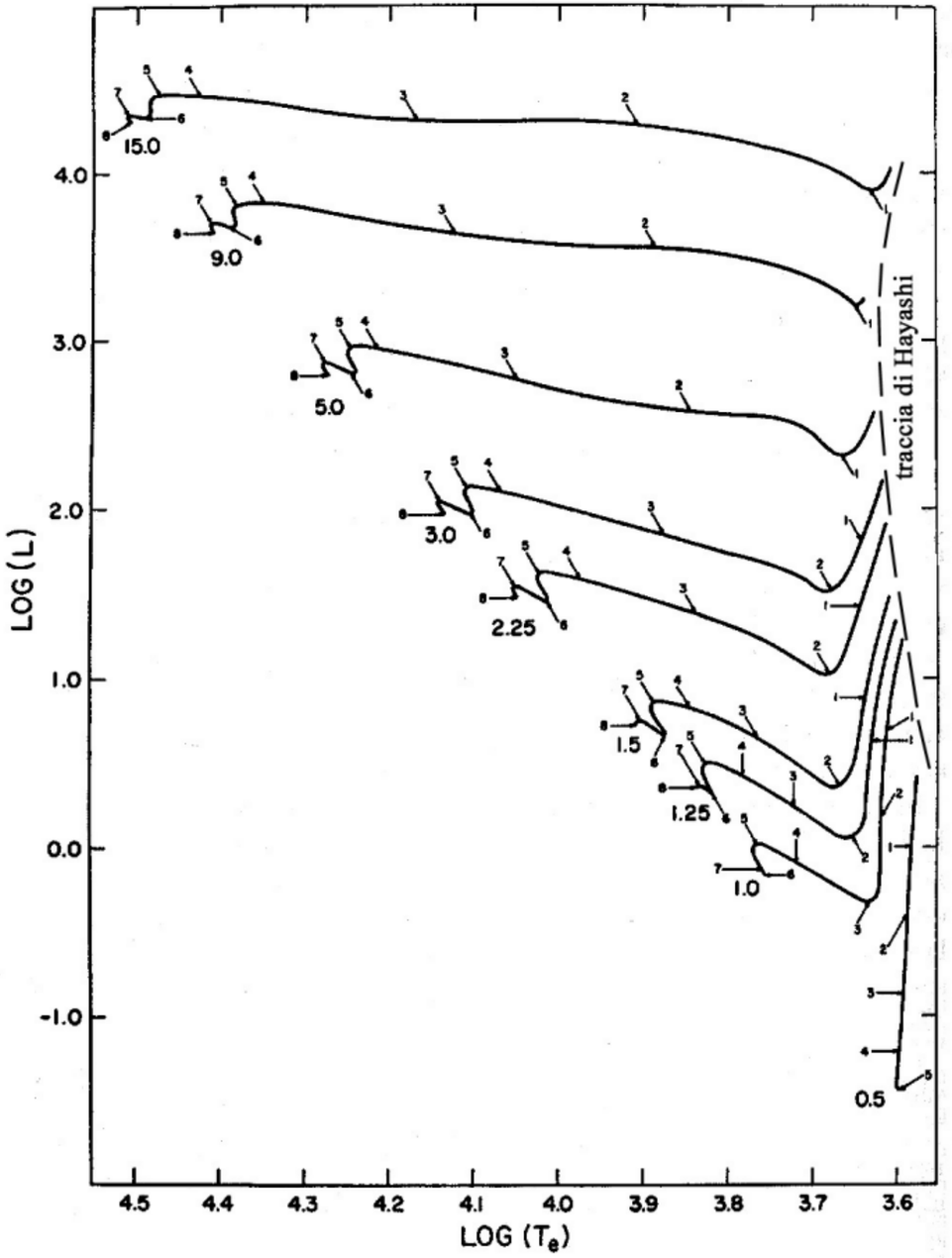
Several stages of hydrostatic equilibrium are reached during the evolution. The first occurs when the temperature reaches about 1800 K, where the dissociation of hydrogen molecules causes a drop in pressure. The core contraction then resumes, and a new hydrostatic equilibrium is established. The process repeats when the contraction raises the temperature to values that allow the ionization of hydrogen ($\approx 10^4$ K) and subsequently of helium ($\approx 10^5$ K). A global hydrostatic equilibrium is finally achieved when the radius becomes 5–10 times that of the present Sun.



Evolutionary tracks of the protostar according to Hayashi. Evolutionary times are measured from the formation of a central core in hydrostatic equilibrium (point A).

Pre-main sequence phase

Once hydrostatic equilibrium is reached, the protostar becomes convective because its opacity is high compared to that of radiation corresponding to a blackbody at a temperature of about 1000 K. Convective energy transport is highly efficient, and therefore the surface of the protostar emits a relatively high luminosity. A thin radiative layer defines the surface temperature, $T_e$ff.



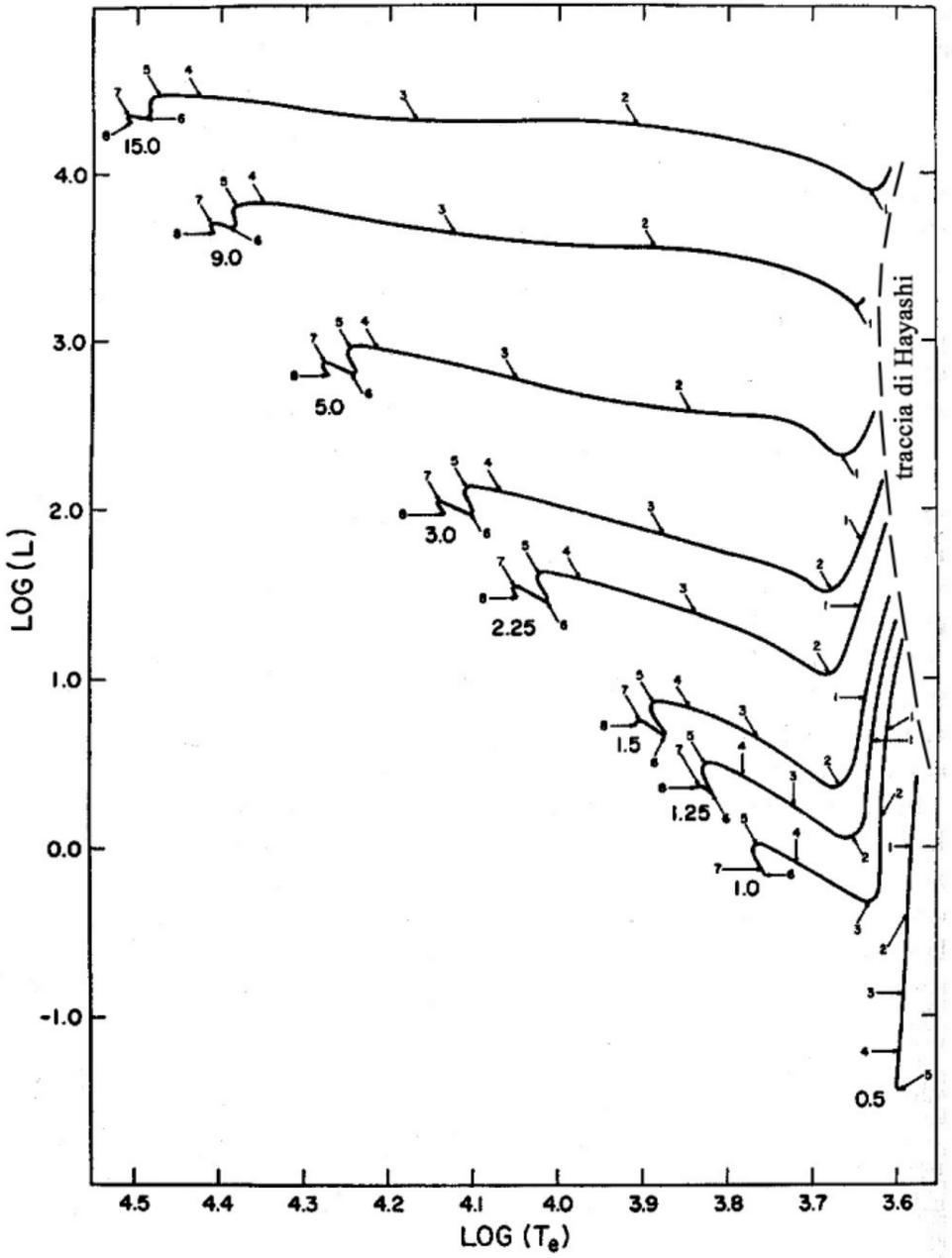
Tracce evolutive teoriche in pre-sequenza

Pre-main sequence phase

Hayashi Convective Phase

Hayashi demonstrated that fully convective stellar models lie along a line in the Hertzsprung–Russell diagram known as the *Hayashi track*. To the right of this line, no equilibrium structures can exist, since energy transport becomes inefficient at low temperatures; consequently, no stable configurations can form, and the structures collapse toward the Hayashi limit on a free-fall timescale. The protostar lies at a point along the Hayashi limit line and evolves according to the Kelvin timescale.

$$t_{Hayashi} = t_K \simeq 10^7 \left(\frac{M}{M_\odot}\right)^2 \left(\frac{R_\odot}{R}\right)^3 \left(\frac{T_{eff,\odot}}{T_{eff}}\right)^4 \text{ anni}$$



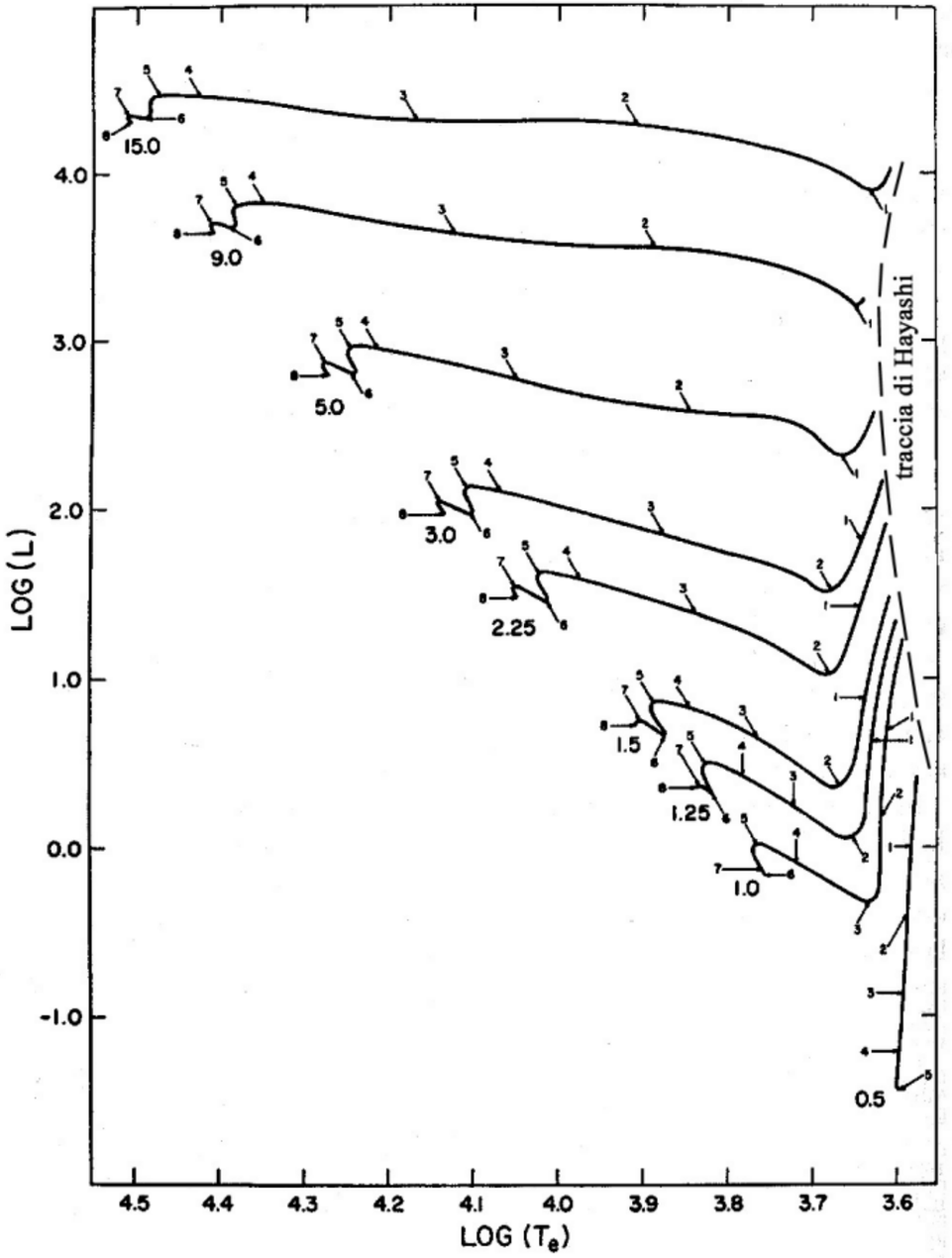
Theoretical Evolutionary tracks

Pre-main sequence phase

Henyey Radiative Phase

The increase in temperature and the ionization of hydrogen and helium lead to a decrease in opacity, allowing radiative energy transport to take over and rapidly extend throughout the protostar. At this stage, the central temperature has risen sufficiently for the first thermonuclear reactions to ignite. The complete CNO cycle is not yet active, but exothermic reactions already occur through the conversion of ^{12}C into ^{14}N.

In addition, energy is produced through deuterium burning in the *p*–*p* chain.

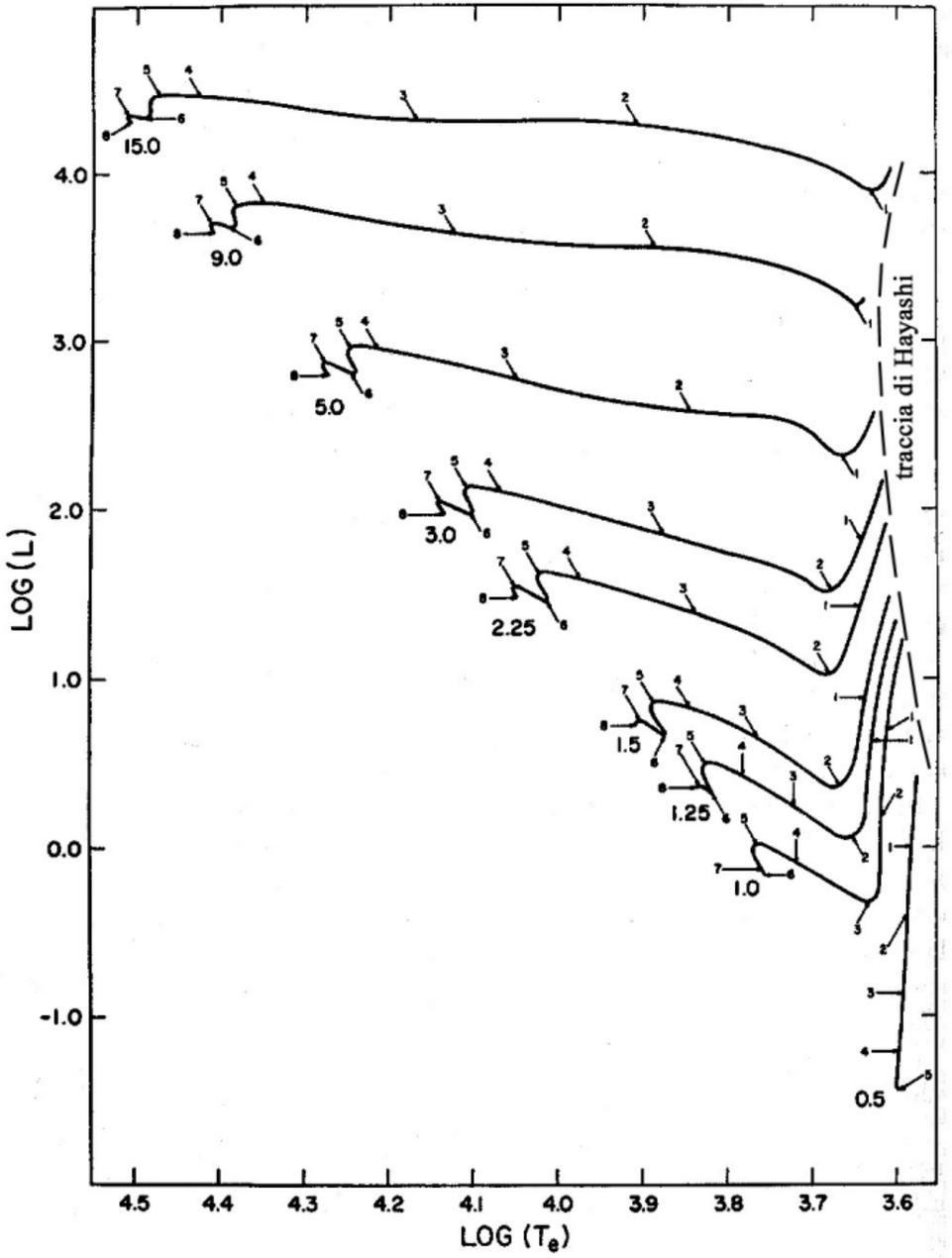


Theoretical Evolutionary tracks

Pre-main sequence phase

Henyey Radiative Phase

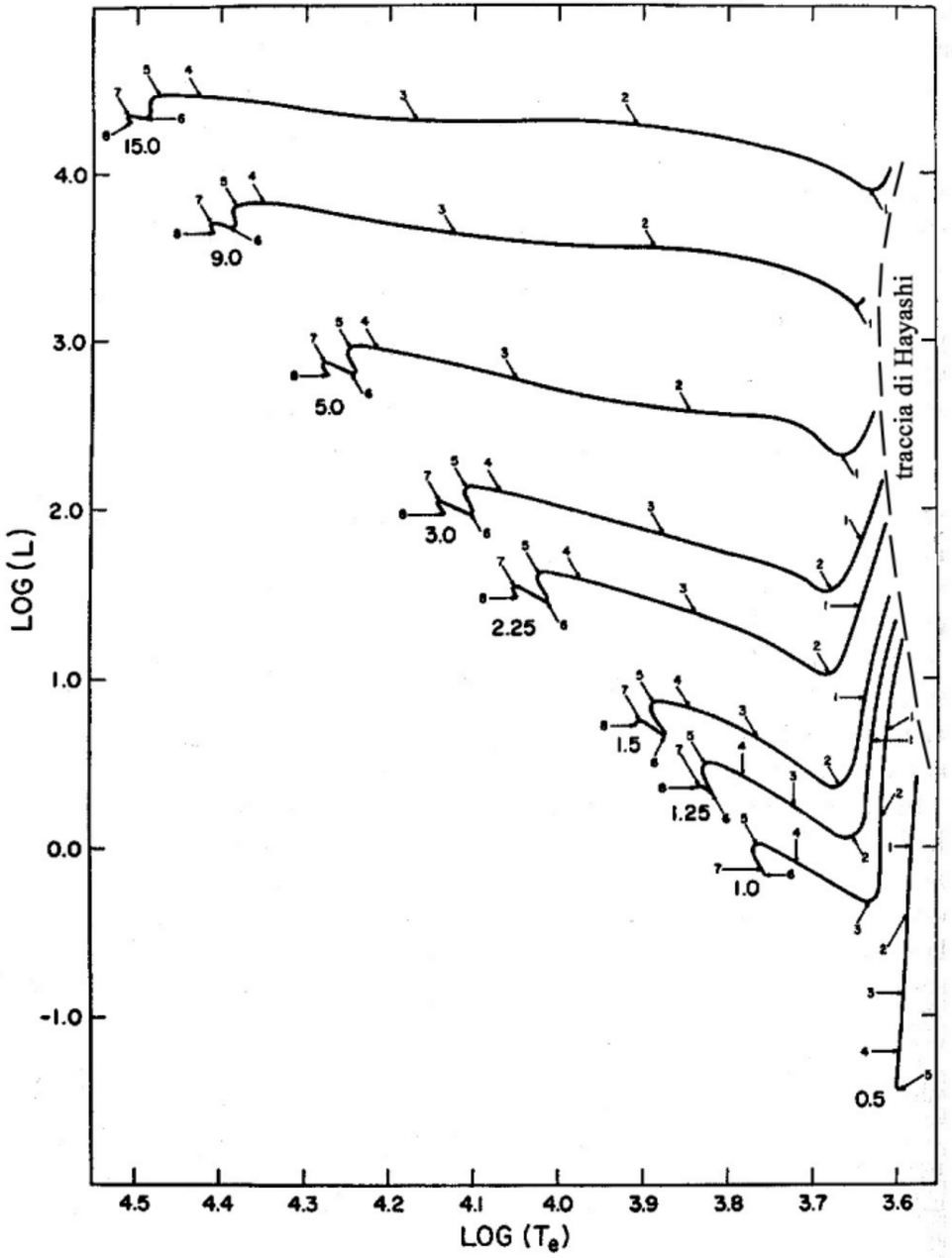
The internal and surface temperatures (Teff) increase, causing the evolutionary track to move to the left on the HR diagram, often accompanied by a slight increase in luminosity (the Henyey track). In higher-mass stars, this transition from convective to radiative occurs very early because they reach higher temperatures quickly, as the gravitational energy released is greater. Lower-mass stars, with masses below that of the Sun, have more difficulty reaching the radiative phase because their temperatures are lower. Stars with a mass ≤ 0.5 M⊙ do not exhibit any radiative phase.



Theoretical Evolutionary tracks

End of the phase of pre-main sequenc

The pre-main-sequence phase ends when the protostar reaches a sufficiently high central temperature to complete the thermonuclear reactions of the p–p chain at $T_c \approx 4 \times 10^6$ K. At this point, the star settles into equilibrium at an almost constant radius, and the energy production from gravitational contraction becomes completely negligible. Only stars within a certain mass range can reach the main-sequence phase. Stars with masses < 0.08 M⊙, due to their low gravity, never attain temperatures high enough to ignite hydrogen burning.



Theoretical Evolutionary tracks

Stellar Evolution in Early Phases of Gravitational Contraction

Chushiro HAYASHI

Department of Nuclear Science, Kyoto University, Kyoto

(Received August 28, 1961)

Abstract

The surface condition for red giant stars worked out in the previous paper indicates that stars lie in the low luminosity and low temperature region of the *H-R* diagram cannot be in equilibrium so that the evolutional path of contracting stars in this region will be different from that calculated by HENYEY et al. The age of these stars along the loci of quasi-static solutions is calculated. The result seems to explain well the *H-R* diagram of a young cluster NGC 2264.

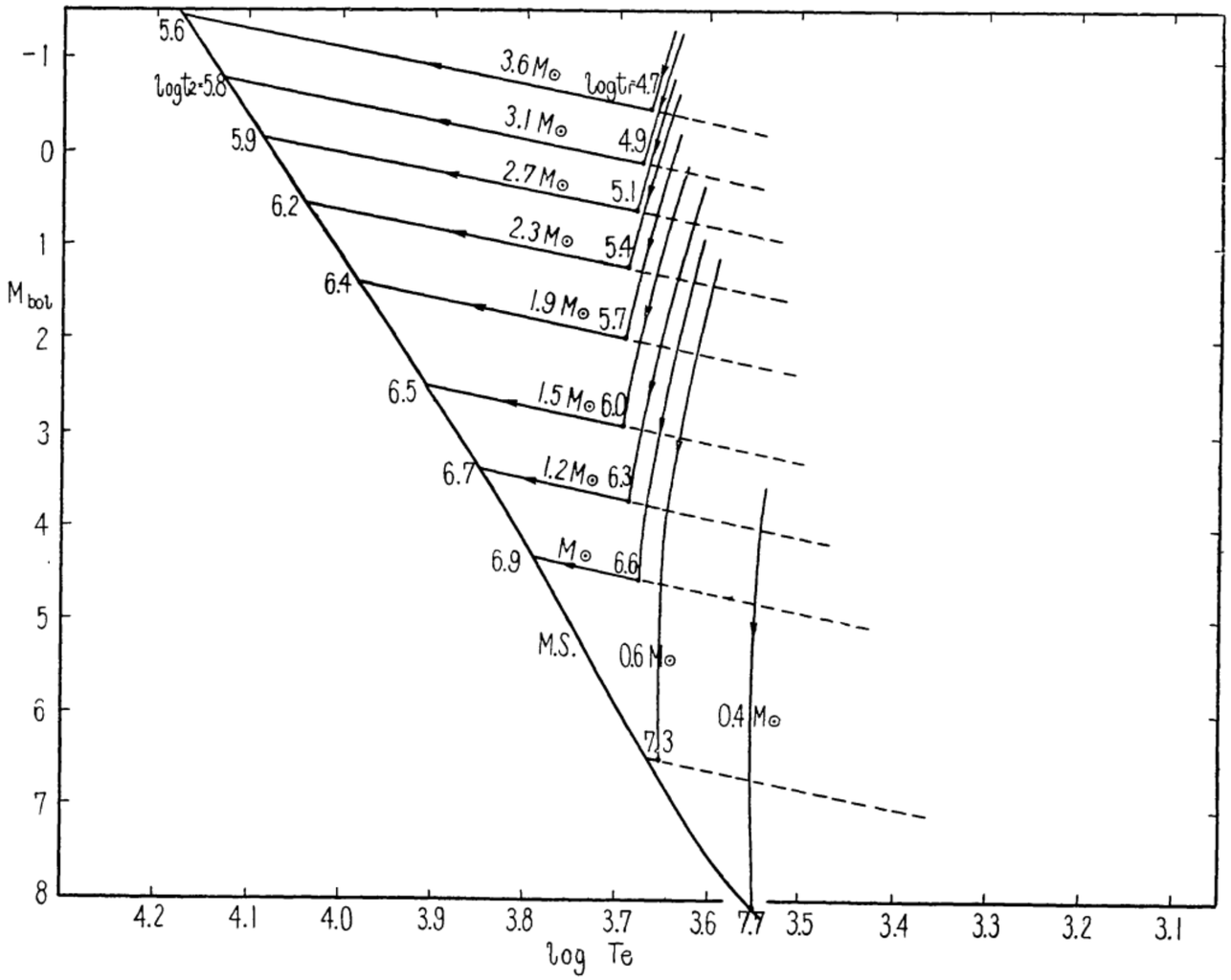
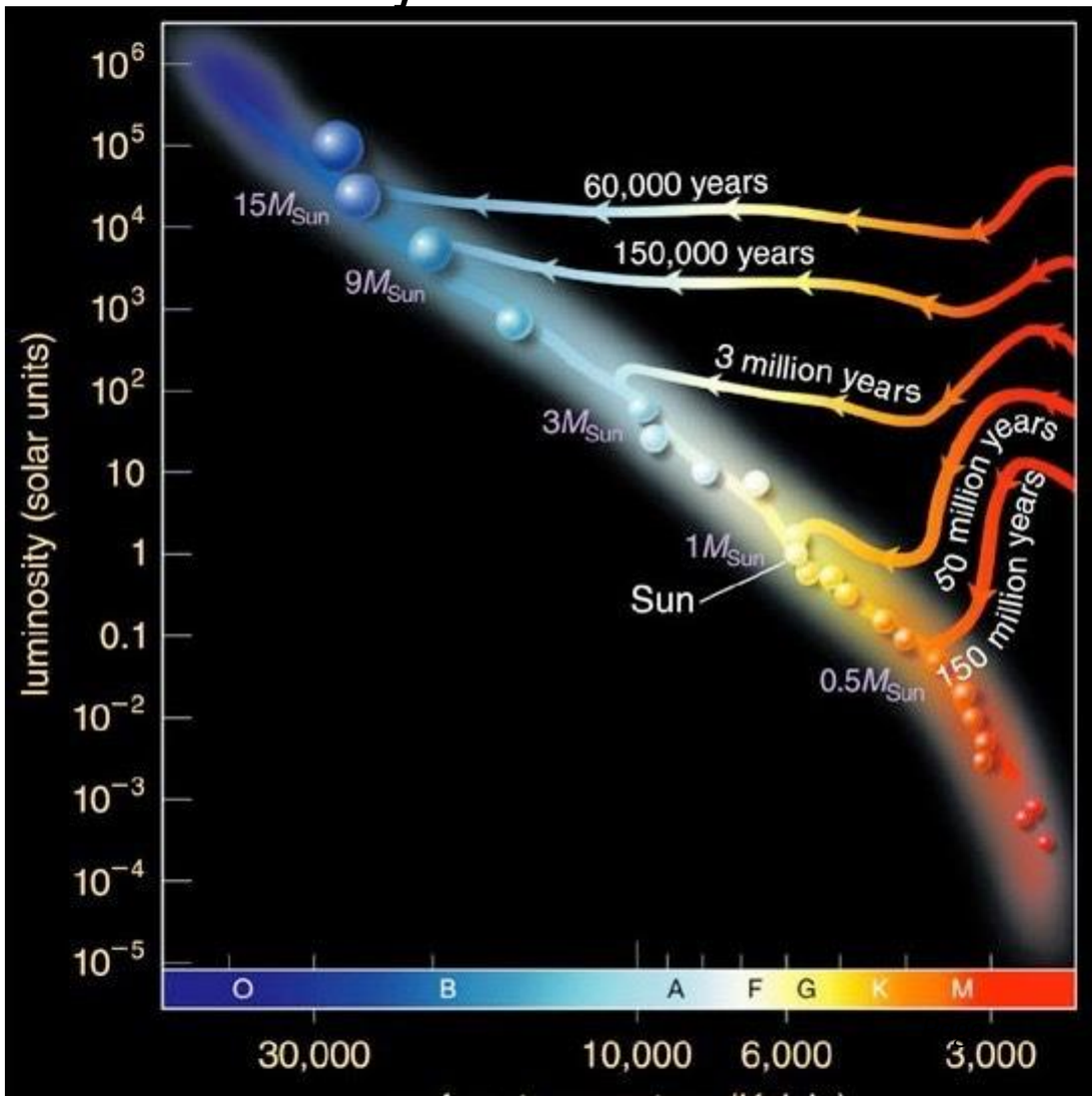


FIG. 2. Evolutional tracks and ages of stars with different masses in gravitational contraction. t_1 and t_2 denote the ages (in yeasr) at the turning point and on the main sequence, respectively.

How long this collapse takes depends on the mass of the forming star. A 15 solar mass protostar may collapse in only 60,000 years while a star half the mass of the Sun would take around 150 million years.

As we will see soon, this is longer than the lifetimes of massive stars, which means that massive stars forming in a cluster can collapse onto the main sequence, complete their hydrogen burning and finish their lives before a low mass star has even made it onto the main sequence.

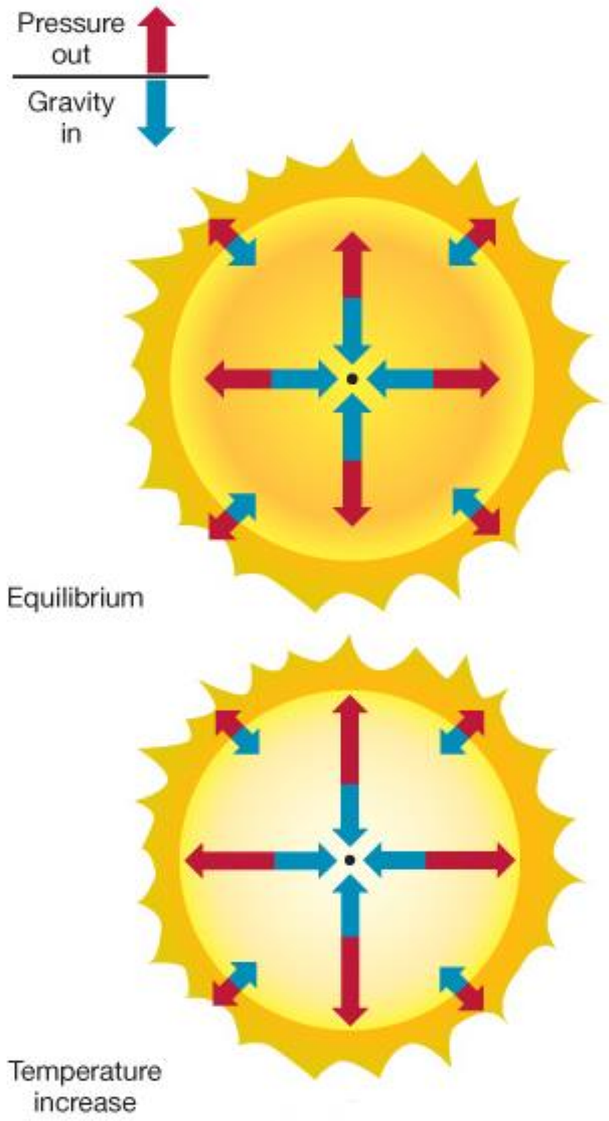


Next subjects

- **Evolution during the Main Sequence**
  Evolution during the main sequence with differences between stars < 1.5 M⊙ and more massive ones.
- **Post-Main-Sequence Evolution**
  Evolution after the main sequence and differences between stars with masses < 2 M⊙ and more massive ones.
- **Final Stages of Stellar Life**
  Final stages of stellar life as a function of their mass.

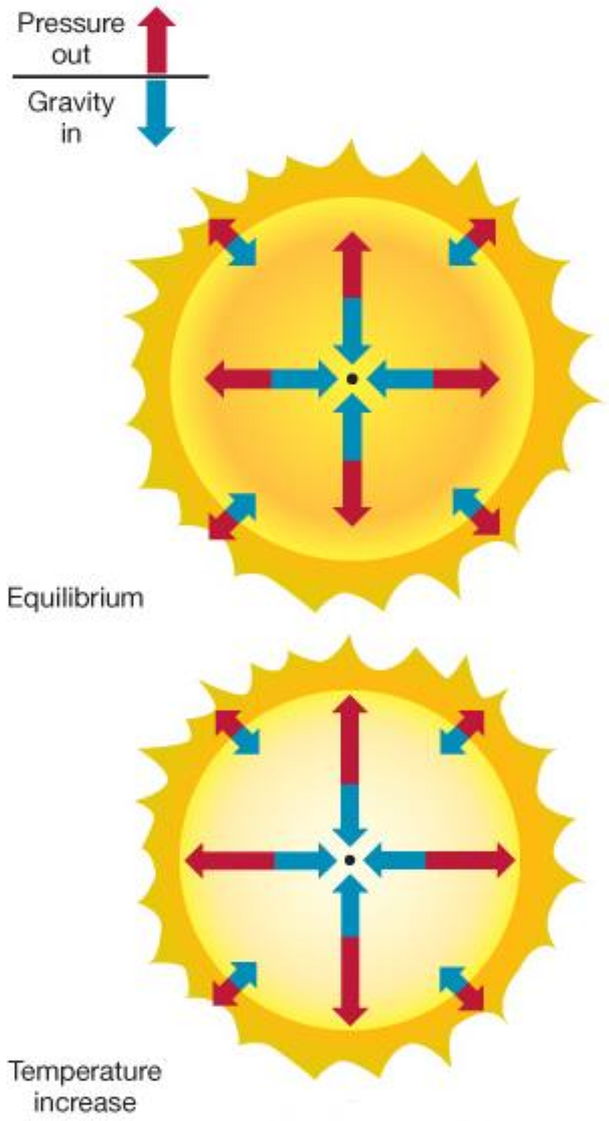
Main Sequence phase

During its stay on the Main Sequence, any fluctuations in a star's condition are quickly restored; the star is in equilibrium



Leaving the Main Sequence

During its stay on the Main Sequence, any fluctuations in a star's condition are quickly restored; the star is in equilibrium:



- A temperature increase in core increases fusion rate

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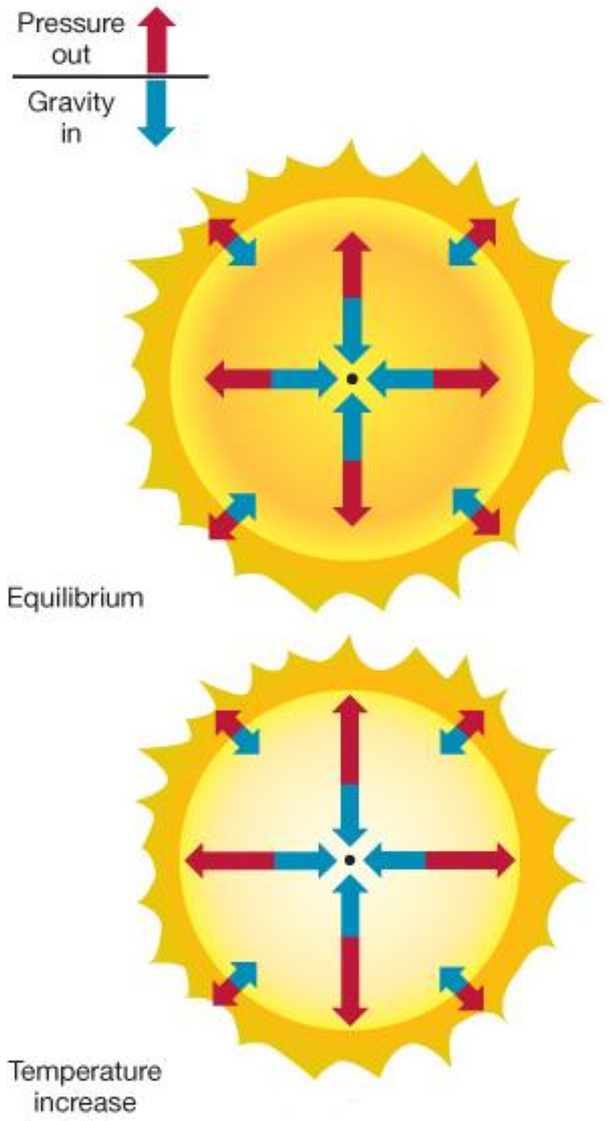

- A temperature increase in core increases fusion rate
- This increases pressure

Leaving the Main Sequence

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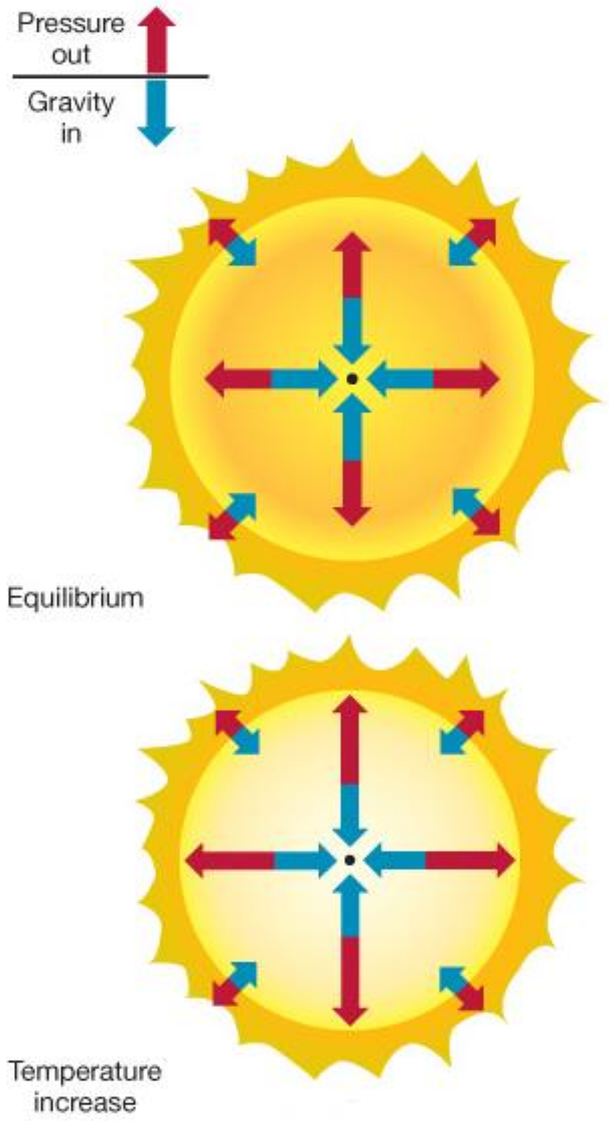

- A temperature increase in core increases fusion rate
- This increases pressure
- Star expands

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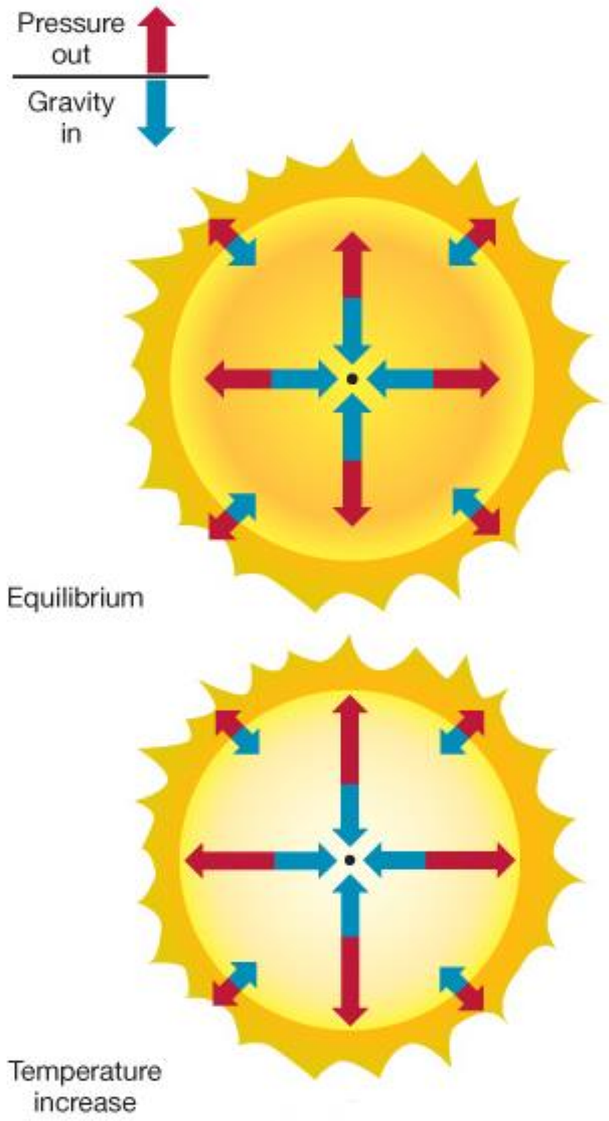

- A temperature increase in core increases fusion rate
- This increases pressure
- Star expands
- Expansion causes cooling

Leaving the Main Sequence

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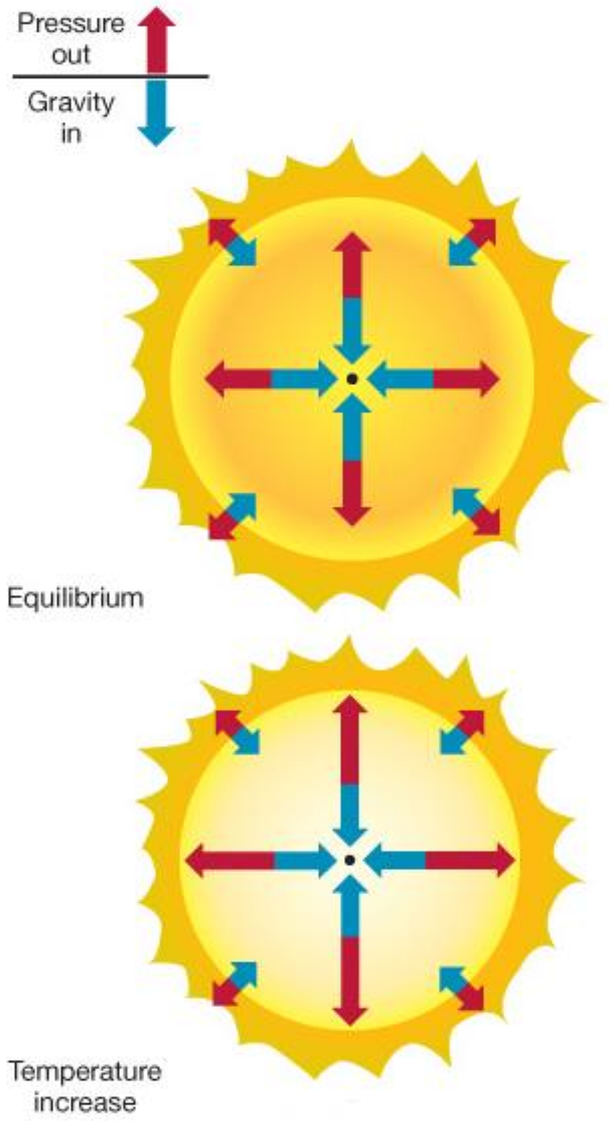

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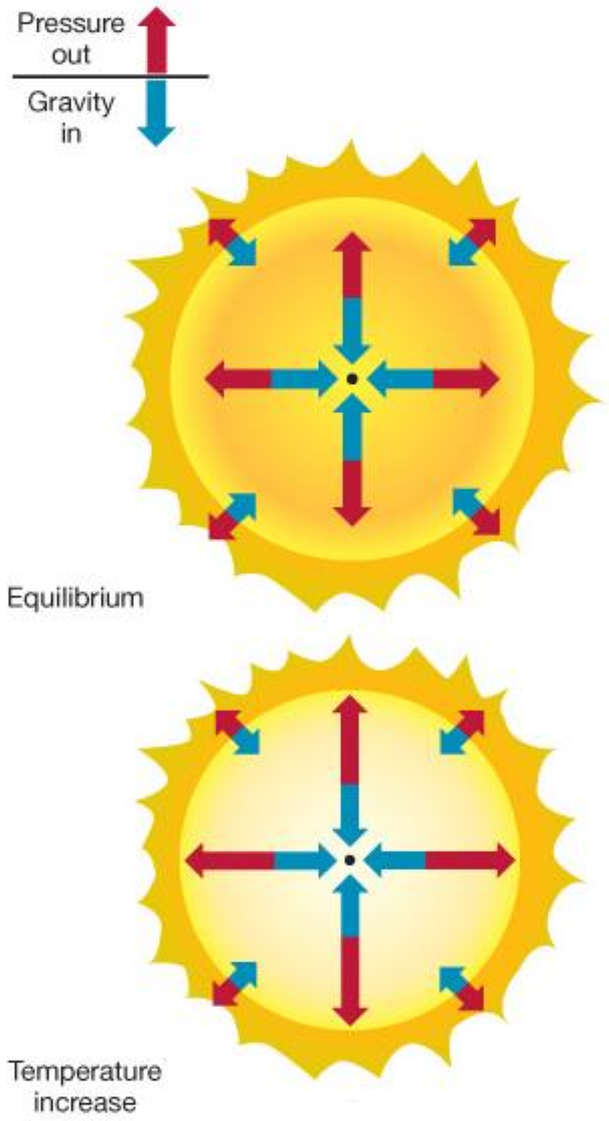

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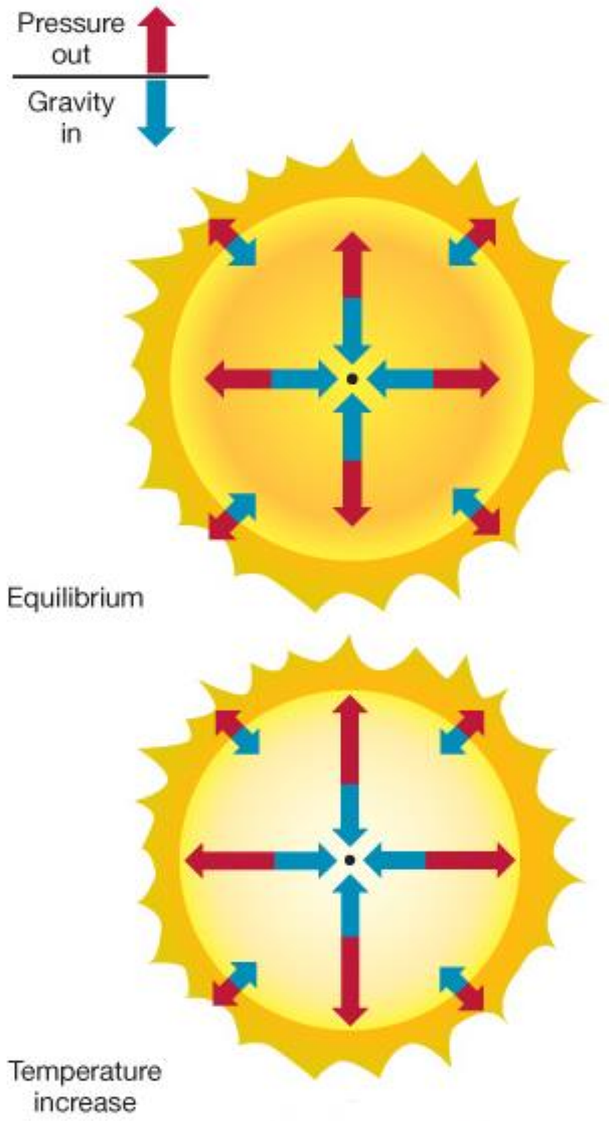

- A temperature increase in core increases fusion rate
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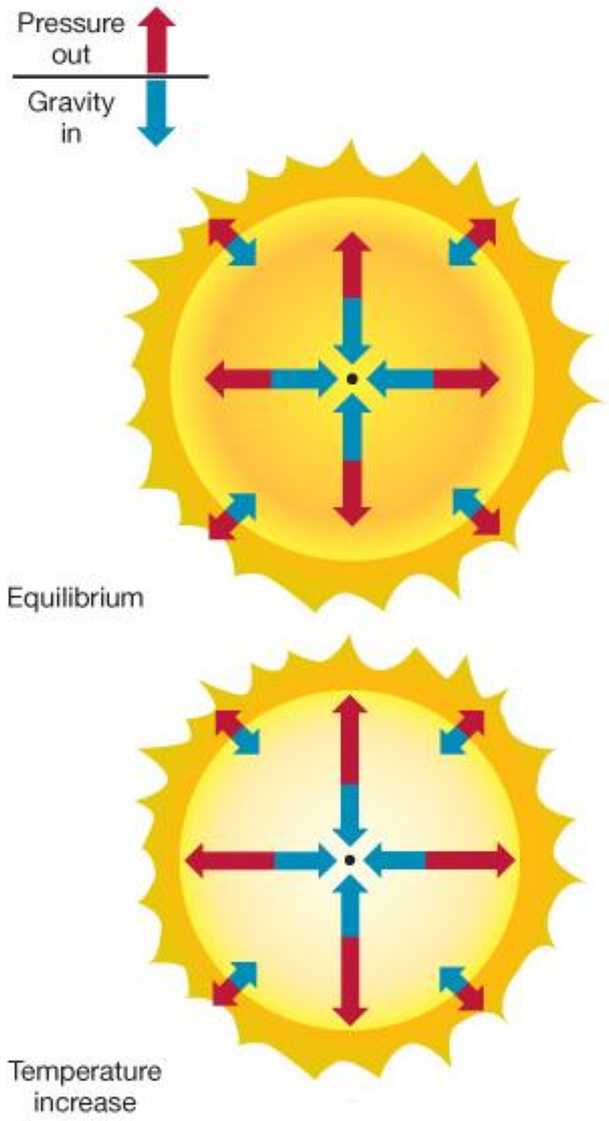


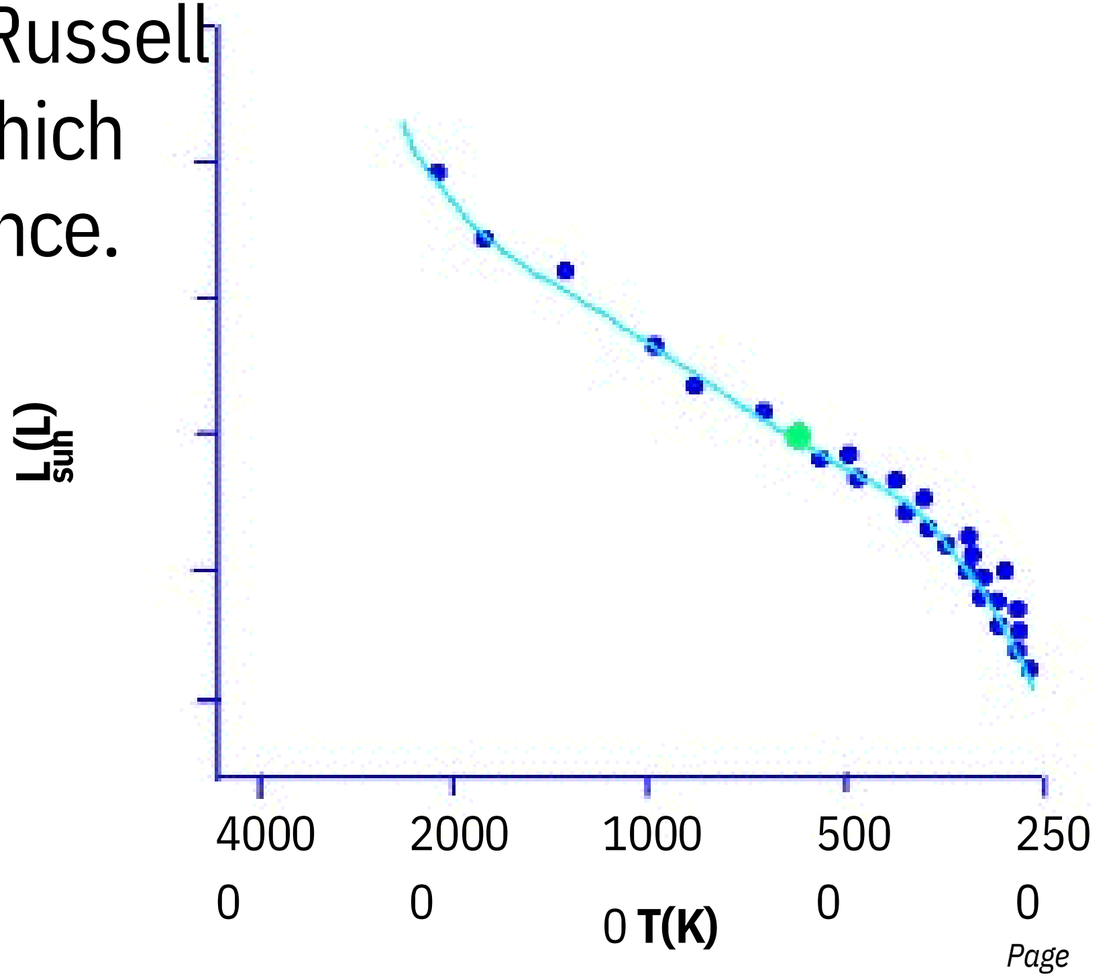
- A temperature increase in core increases fusion rate
- This increases pressure
- Star expands
- Expansion causes cooling
- Fusion rate falls
- Pressure decreases
- Star returns to original size
- A temperature decrease causes the opposite

The fact that gravity determines the structure of the star explains why it is the *mass* which is the most important variable in what it looks like.

This is exemplified in the Herztsprung-Russell diagram, the most obvious feature of which – is a *mass* sequence, not a *time* sequence.

Homologous models told us this!

An F-type star formed as an F-type star, and will remain an F-type star during entire main sequence stage.

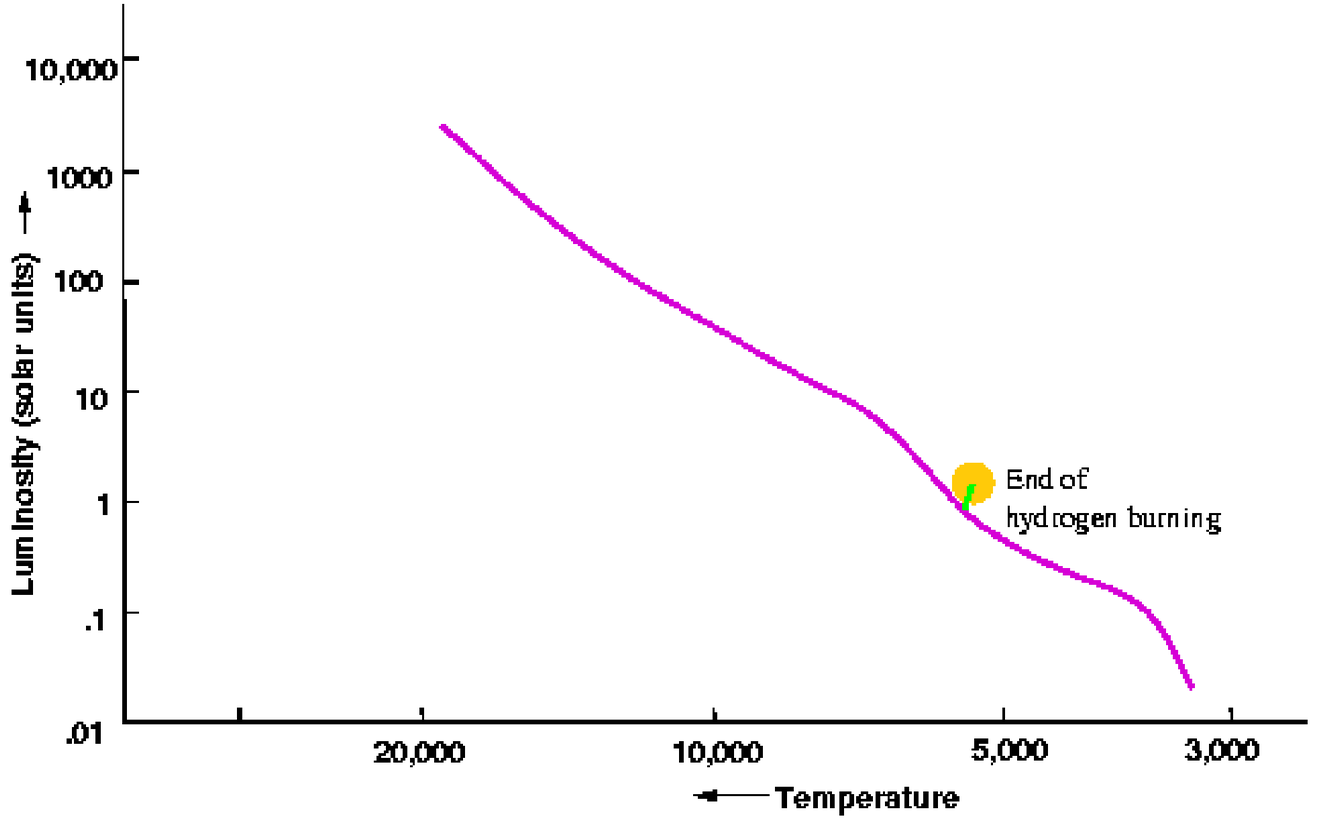


Changes on the main sequence

Stars do change somewhat while they are on the main sequence. While a star is on the main sequence, it is burning hydrogen in its core. Helium is formed, and gradually builds up as a sort of ash: *helium poisoning*.

Helium can't burn, because it needs a temperature of about 100 million degrees before it can ignite. So, it just builds up in the core.

As each fusion converts four H atoms to one He atom, the core of the star has fewer atoms in it, so the pressure goes down. Gravity squeezes the core more tightly, which increase the temperature, which increase the rate of fusion. This produces more energy, which makes the outside layers of the star expand a bit, which makes the star a bit brighter and a bit cooler.



This occurs differently depending on the mass of the star.

Stars with 0.08 < M < 0.26 M⊙

Stars of this mass have a very simple evolution: they remain fully convective, which means that all their hydrogen is available for burning. As a result, they live for a very long time, longer than the Hubble time.

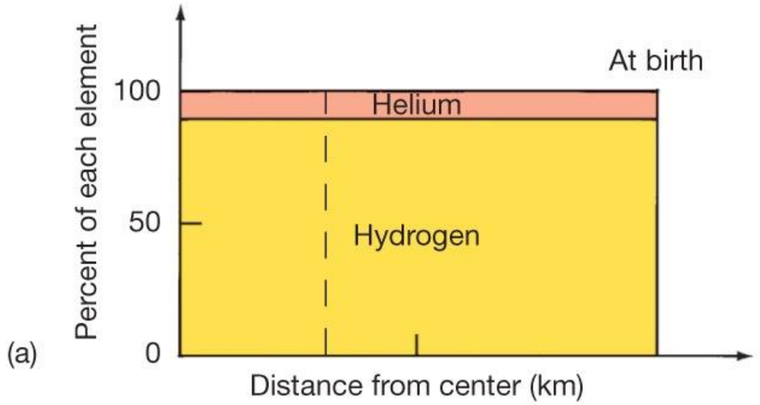
Stars with M < 0.08 M⊙

They become degenerate before they are able to ignite hydrogen and are known as brown dwarfs.

Stars with M < 1.5 M⊙

They reach $T_c < 1.8 \times 10^7$ K and burn hydrogen through the p–p chain. This chain produces little energy and has a weak temperature dependence; the nuclear cores are radiative and extend over a large portion of the stellar interior. The opacity of the outer regions of the star is high due to the low temperature of the stellar gas, and the stellar envelope becomes convectively unstable.

Stars of masses 0.26Msun<M<1.5Msun



Even while on the Main Sequence, the composition of a star's core is changing:

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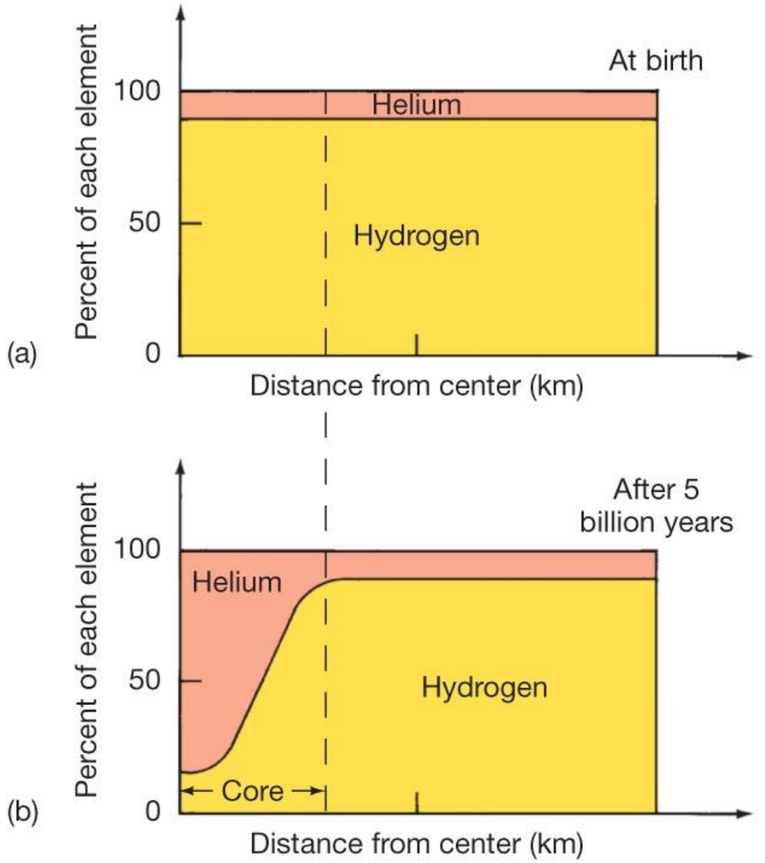

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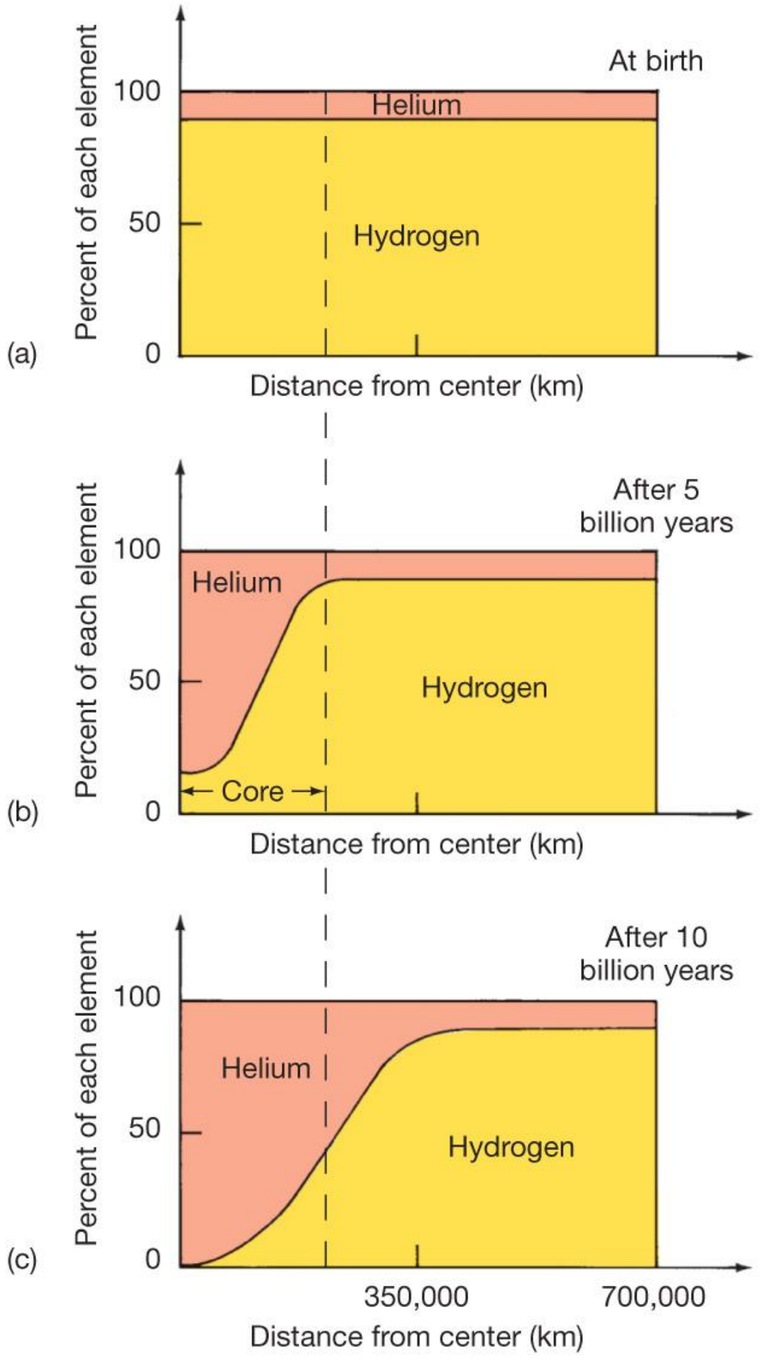

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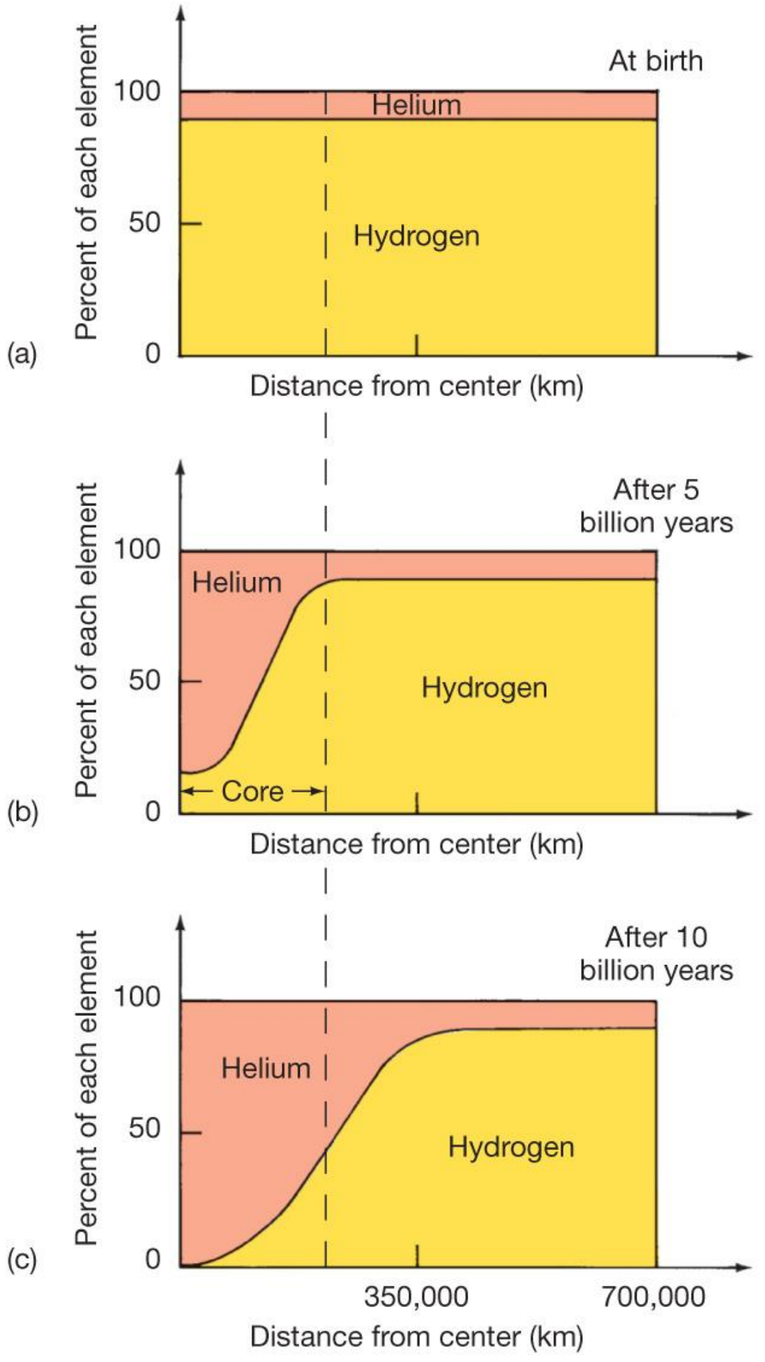
Even while on the Main Sequence, the composition of a star's core is changing:

As the fuel in the core is used up, the core contracts



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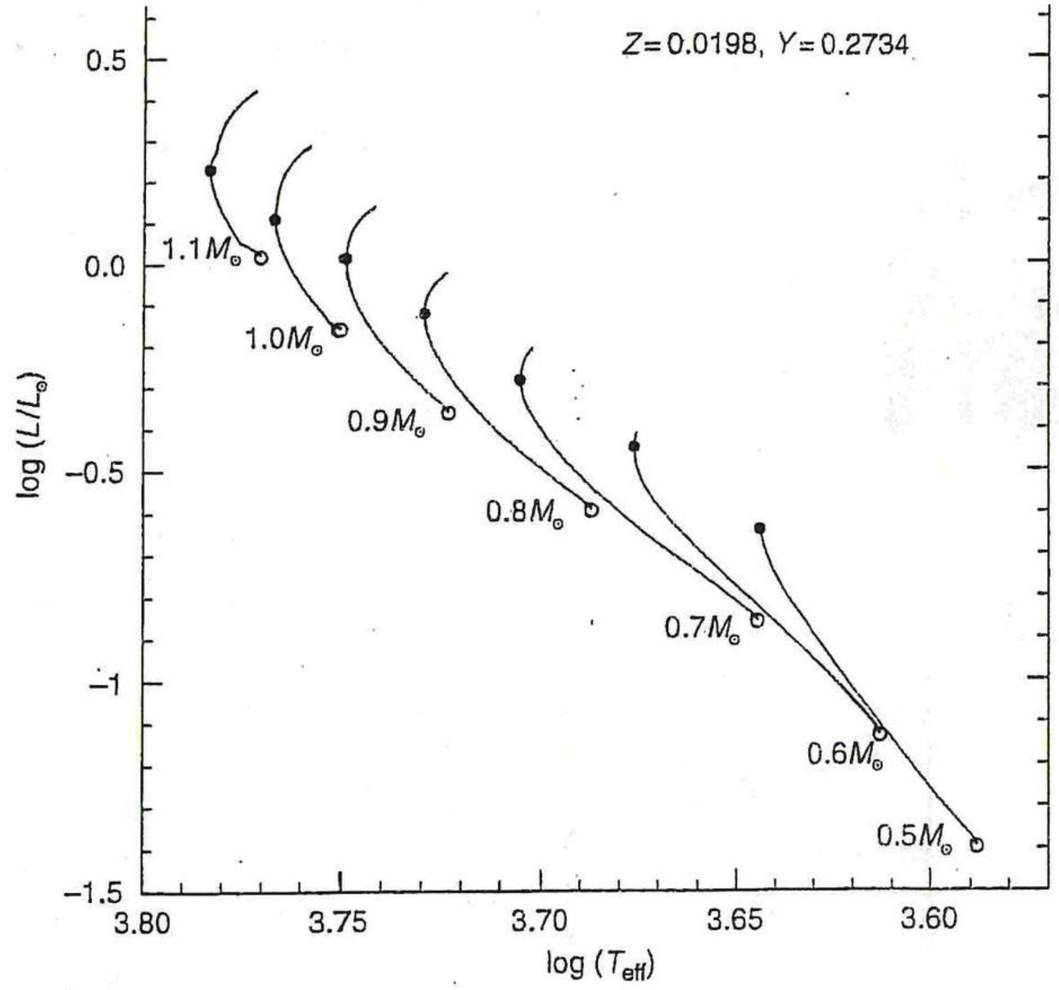
Evolution during H-burning (theoretical tracks)



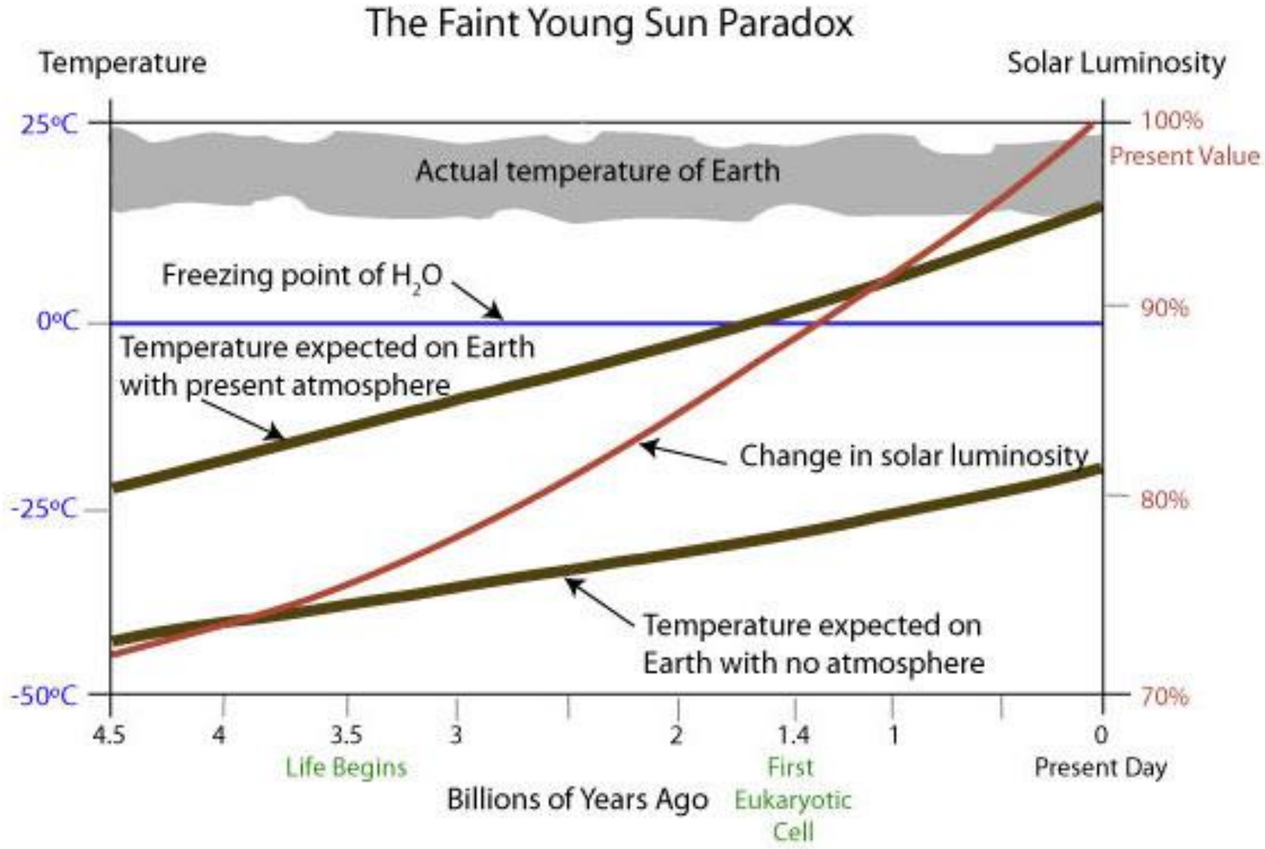
SALARIS & CASSISI (WILEY)

Figure 5.3 The HRD for low-mass stars of different mass during the core H-burning phase. Th solid dot marks the location of the turn off along each track

EXTRA

These changes in the Sun will have a profound effect on the Earth. When the Sun began its main sequence life about 5 billion years ago, it was only 70% as bright as it is now. In another 5 billion years, it will be roughly twice as bright, which will raise the average temperature of the Earth at least 19° C.

In fact, we have a hard time reconciling these predicted temperatures with the geological record (the “faint young sun paradox”).



Even though the Sun was about 30% dimmer than it is now, the temperature on Earth has been more or less stable.

Stars with masses M>1.5Msun

They reach high central temperatures, $T_c \geq 1.8 \times 10^7$ K, and burn hydrogen via the CNO cycle, which is much more efficient, and therefore have a convective core.
The outer region (envelope), where no energy is produced, is instead radiative. The hydrogen abundance is strongly discontinuous: in the core, convection mixes the material and ensures a homogeneous composition; in the envelope, the original chemical composition is preserved.

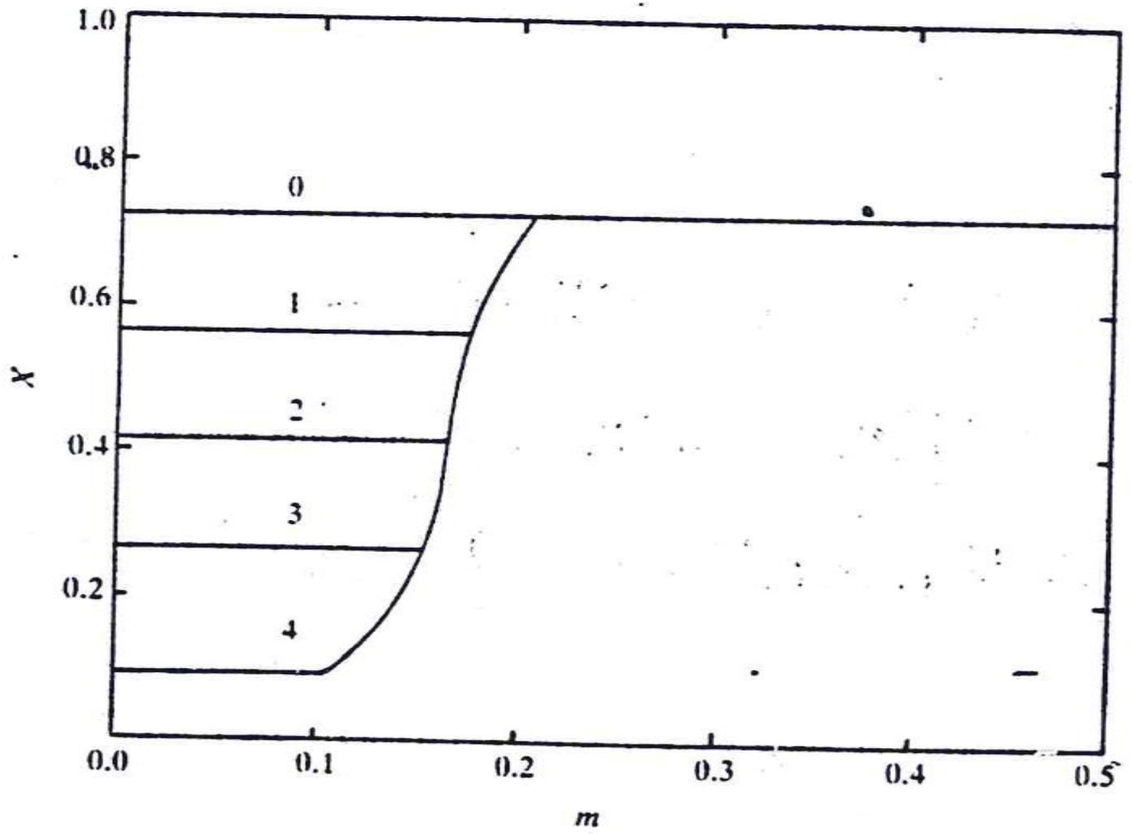


Fig. 63. The depletion of hydrogen in a high mass star. The numbers denote successive stages in the star's evolution.

Evolution durin H burning for a 5Msun star

During evolution, the mass of the convective core decreases as hydrogen is depleted, while the luminosity increases slightly. When hydrogen in the core is exhausted, the entire star contracts rapidly, releasing gravitational energy that increases the luminosity. As a result, the temperature must rise, leading to hydrogen burning in a shell surrounding the core. The star's temperature also increases to compensate for the decreasing radius, causing the evolutionary track to move upward and to the left.

SALARIS & CASSISI (WILEY)

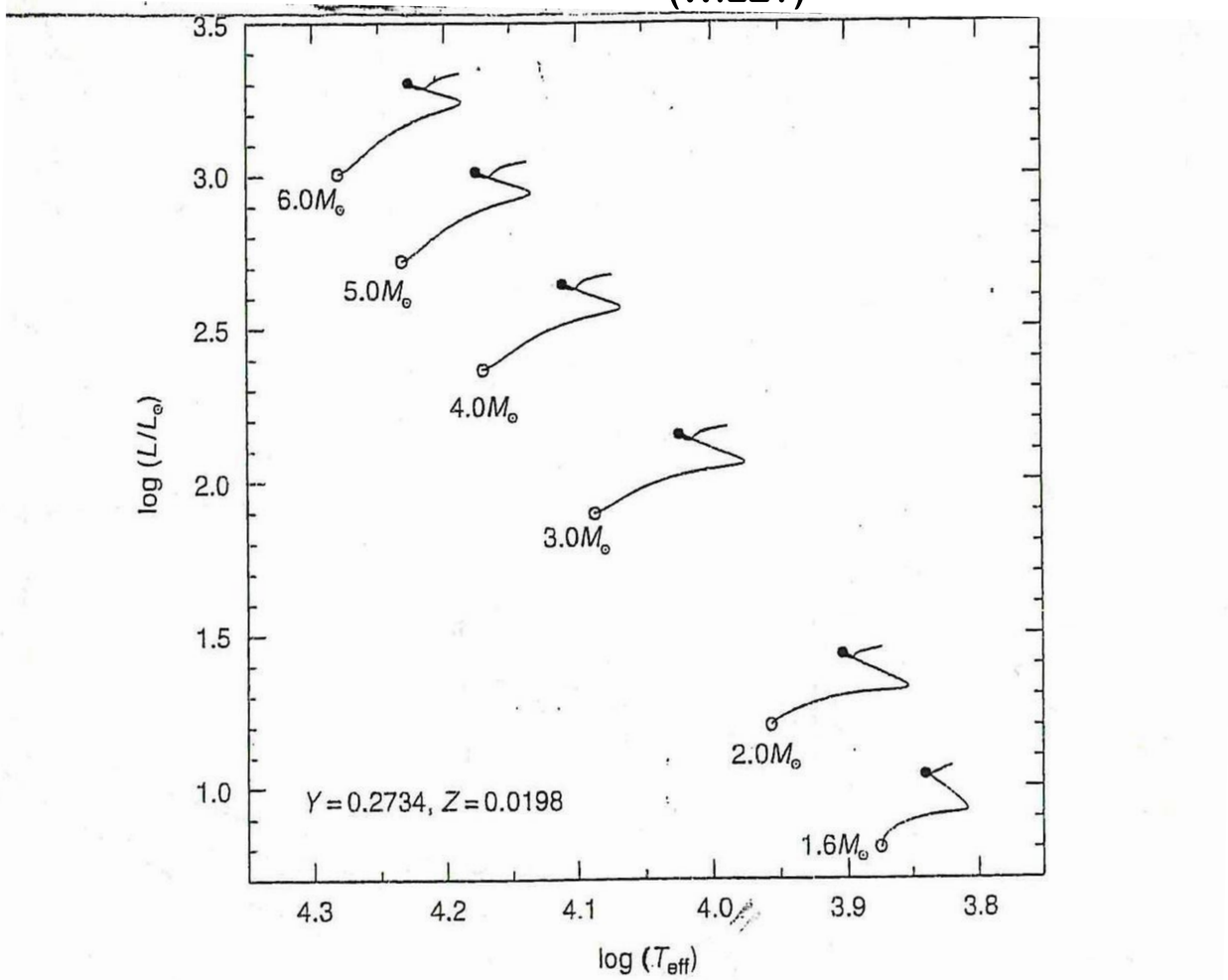
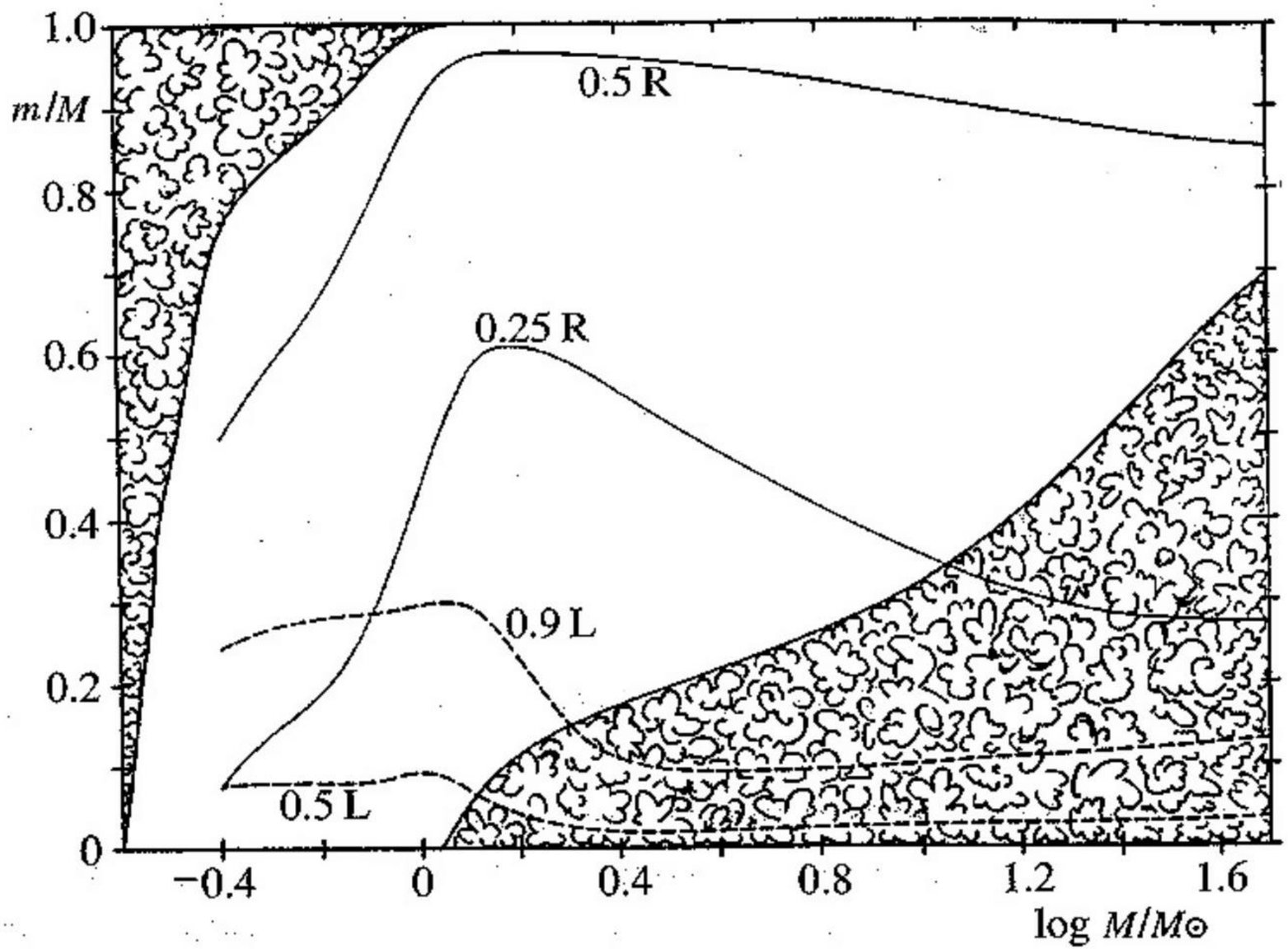


Figure 5.5 Evolutionary tracks of different intermediate-mass stars during the core H-burning phase. The solid dot marks the evolutionary stage equivalent to the turn off point in less massive stars

Convective and radiative zone as a function of the mass

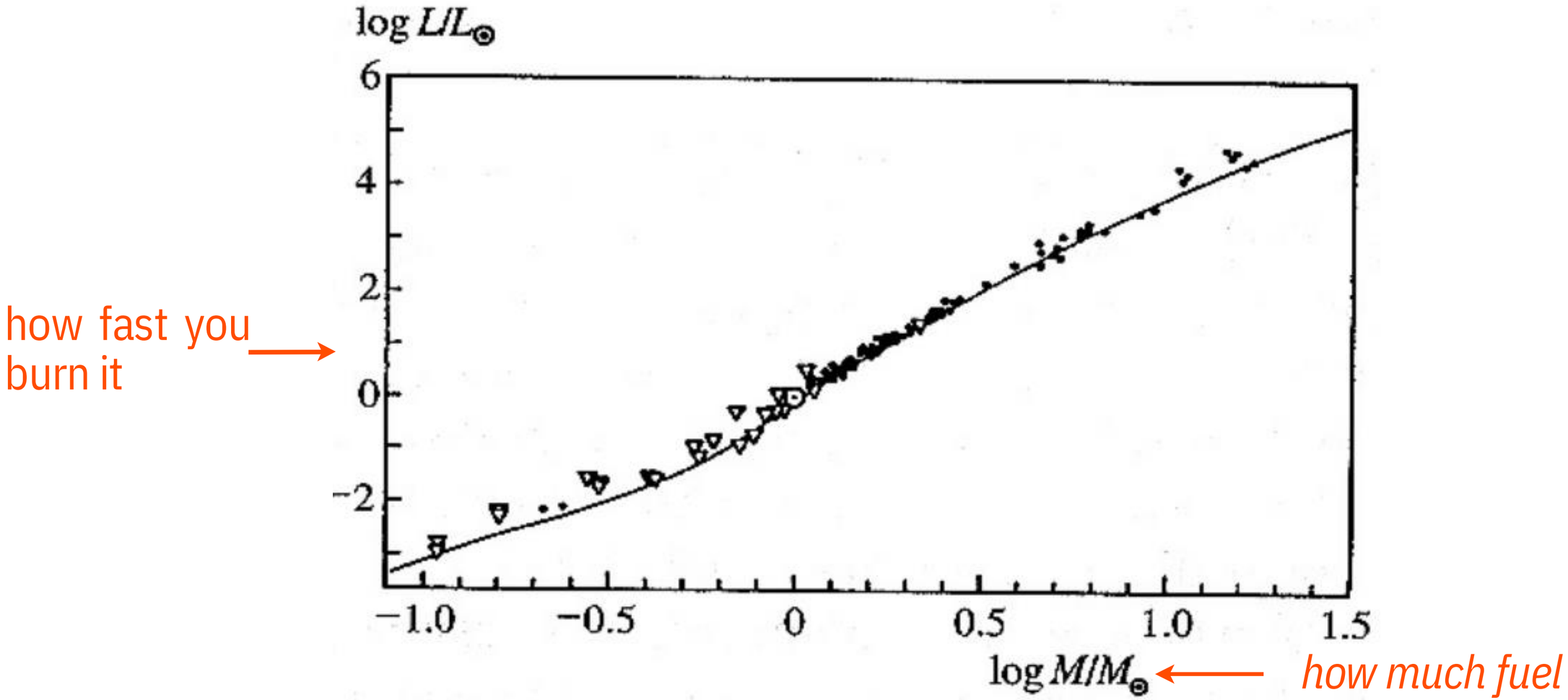


How long a star can burn hydrogen on the main sequence depends on two things:

- how much hydrogen it has; and
- how fast it burns it

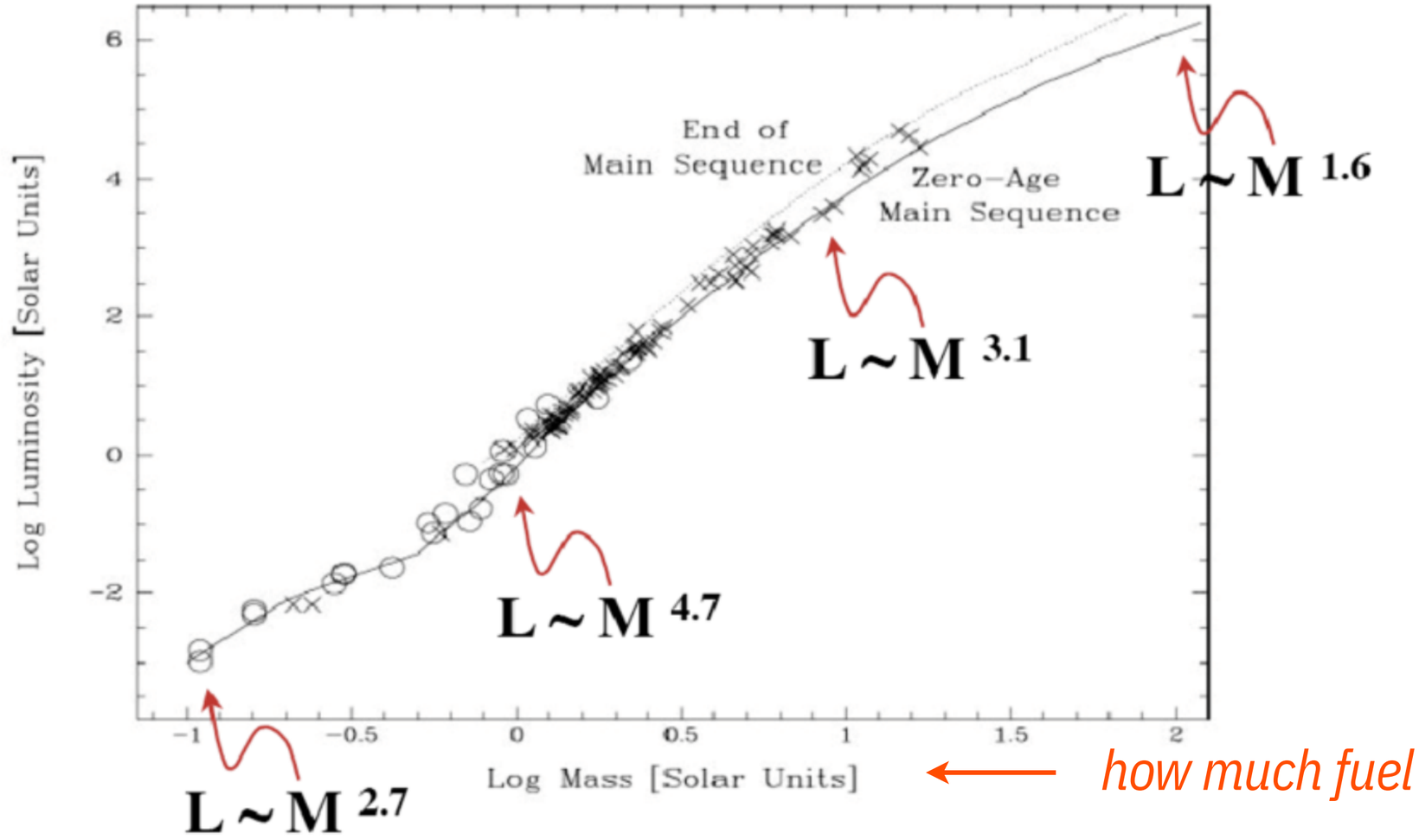
The first is just the star's mass, and the second the star's luminosity.

But we know how those two are related: we derived the mass-luminosity



Here it is again: as we increase the mass of the star, the luminosity increases *enormously*: a factor of 10 increase in mass corresponds to a factor of 3000 increase in luminosity.

how fast you burn it →



← *how much fuel*

Same plot, but more recent from Alksnis 2017

We can write this dependence of the star's luminosity on its mass as

$L \propto M^4$ for stars of $0.4M_{sun} < M < 10M_{sun}$

$L \propto M^b$ for stars of $M < 0.4M_{sun}$ or $M > 10M_{sun}$ with $b < 4$

the lifetime of a star goes like

$t \propto M/L \propto M/M^4 = M^{-3}$

In other words, a factor of 10 increase in mass corresponds to a *decrease* in the lifetime of the star by a factor of 1000.

$$t_{MS} \simeq 0.0006 \frac{Mc^2}{\mathcal{L}} \simeq 10^{10} \left(\frac{M}{M_\odot} \right)^{-3} \text{ anni} \quad .$$

So, we can make a table showing the lifetime for stars of different masses:

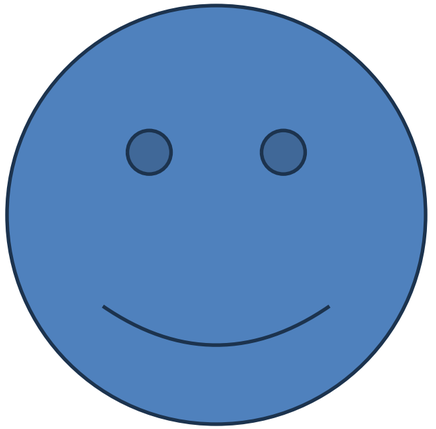
Spectral type	Mass (Sun=1)	Luminosity	Years on main sequence
O5	40	405,000	1×10^6
B0	15	13,000	11×10^6
A0	3.5	80	440×10^6
F0	1.7	5.4	3×10^9
G0	1.1	1.4	8×10^9
K0	0.8	0.46	17×10^9
M0	0.5	0.08	56×10^9

Interestingly, the amount of energy released per kilogram is almost identical for all types of stars.

$$\text{energy per kilogram} = L\,t / M$$

$$\propto M^{3.5} \times M^{-2.5}/M$$

$$= \text{constant}$$



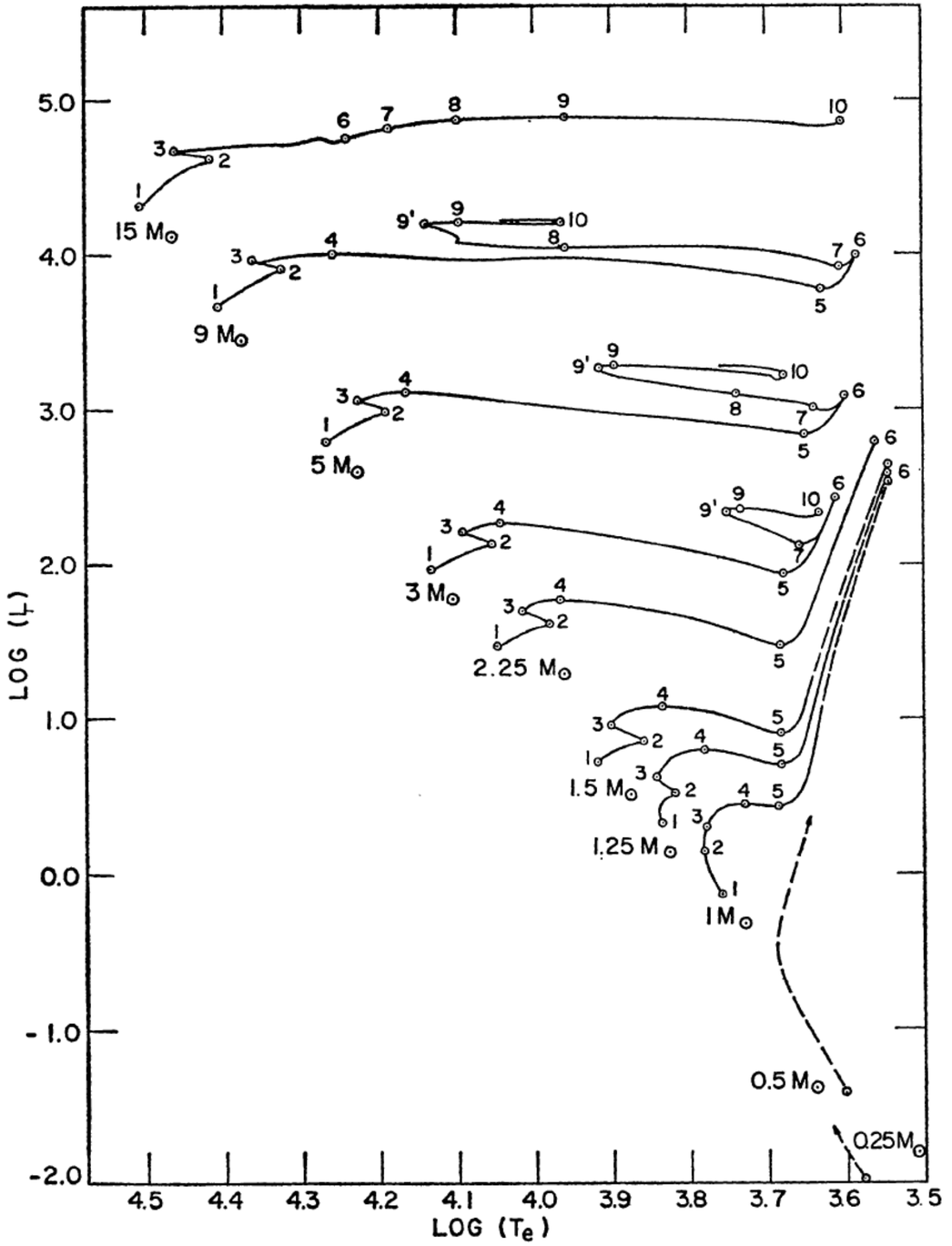
Massive stars are much more luminous, but live for much less time, so the amount of energy extracted per kilogram is identical to the faintest M stars.

This is, of course, because all stars are getting their energy from the same source: nuclear reactions.

Beyond the main sequence

In the figure, the post–main-sequence evolutionary tracks calculated by Iben (1967, *ARA&A*) are shown.
Despite their apparent similarity, the evolutionary histories of low-mass and high-mass stars differ significantly.

It is also important to note that these tracks may vary from author to author and may change over time — some examples will be shown later.



Beyond the main sequence

Stars of mass M>2Msun (as example 5 Msun)

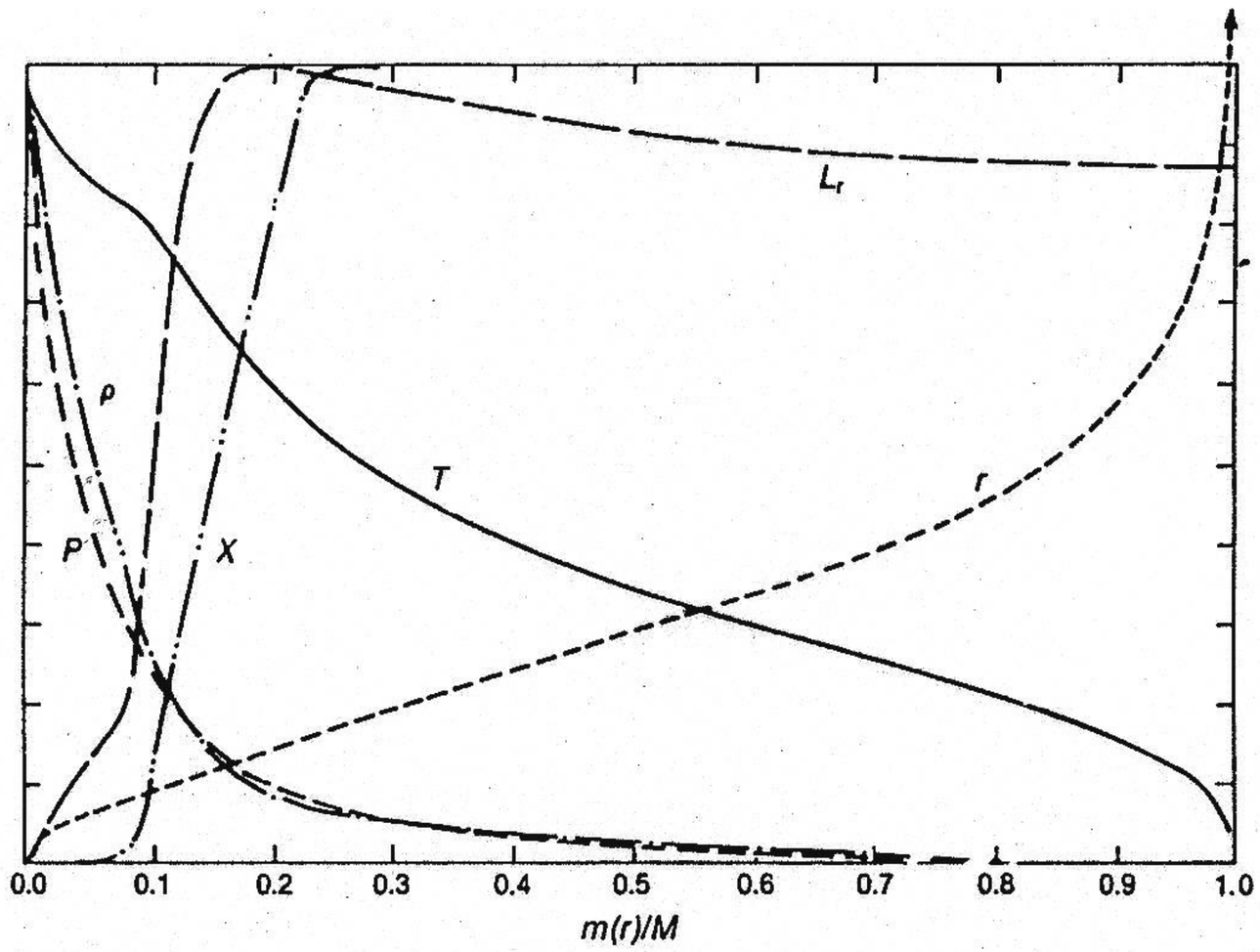
A mass of approx 2 M⊙ corresponds to the threshold above which a red giant ignites helium in a non-degenerate core.

As the hydrogen in the core is depleted, the inner regions of the star begin to contract rapidly and heat up, as predicted by the virial theorem. When hydrogen is nearly exhausted at the centre, the increasing temperature causes a surrounding layer (the circum-nuclear shell) to reach the conditions necessary for hydrogen burning. The sudden ignition of this shell produces an expansion of the stellar envelope.

Stars of 5Msun

The hydrogen-burning shell causes the helium core to increase in size and to become nearly isothermal.

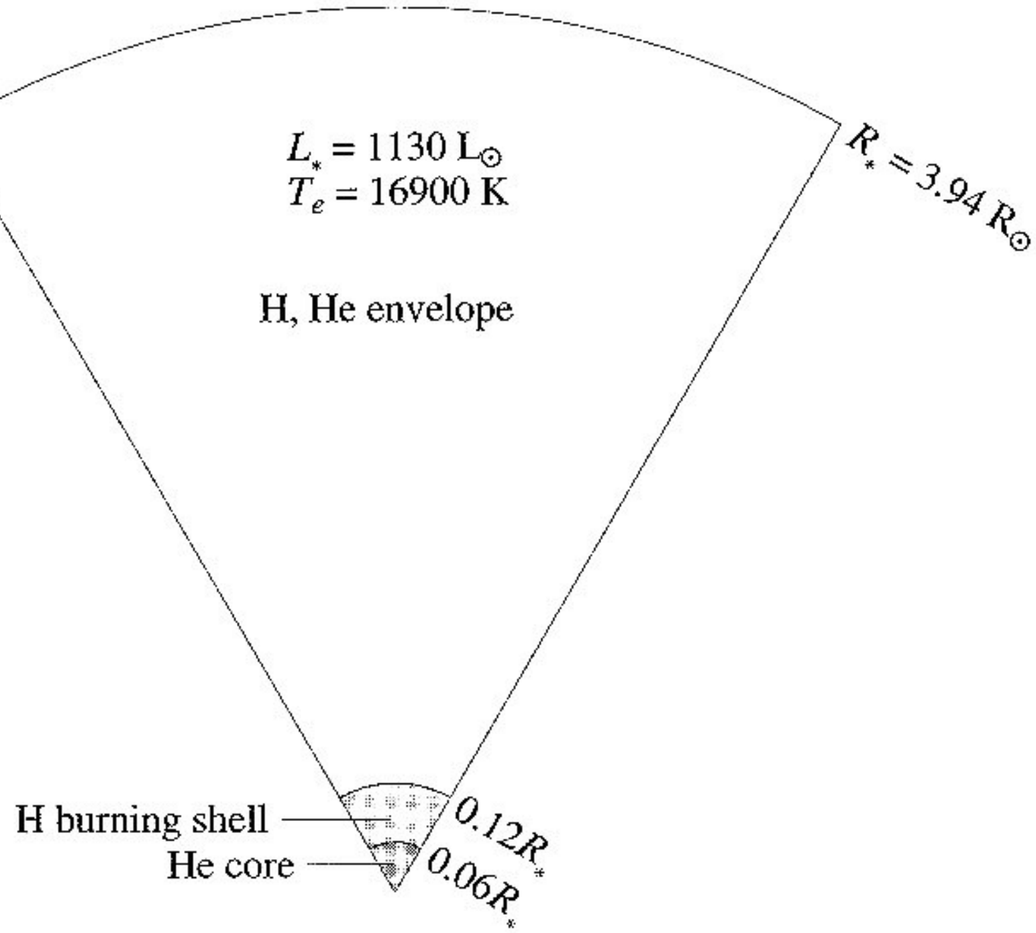
In fact, since no energy is produced in this region (L(r) = 0), from the definition of the radiative temperature gradient it follows that **dT/dr = 0**.



Model of a 5 $M_\odot$ star just after it leaves the main sequence at a time of 6.84461×10^7 years. Maximum value of the ordinate for each curve; $r = 2.9198\ R_\odot$; $P_c = 2.427 \times 10^{17}$ dynes cm^{-2}; $\rho_c = 106.59$ g cm^{-3}; $T_c = 3.6164 \times 10^7$ K; $L = 1.291 \times 10^3\ L_\odot$; and $X_c = 0.708$. The total radius is $R = 3.943\ R_\odot$.

Stars of 5Msun

The stellar structure therefore differs from that of the main sequence: the central core becomes very compact and radiative, while the outer envelope becomes very extended and convective.
The representative point of the star in the HR diagram moves rapidly toward the region of low temperatures and high luminosities, approaching the Hayashi track.



The following principle appears to be generally valid, and provides a way of interpreting the results of detailed numerical calculations:

- Whenever a star has an active shell-burning source, the burning shell acts as a "mirror" between the core and envelope.

$$\text{core contraction} \Longrightarrow \text{envelope expansion}$$

$$\text{core expansion} \Longrightarrow \text{envelope contraction}$$

- The result may be understood by the following argument:
  - To maintain thermal equilibrium, the burning shell must remain at approximately constant temperature due to the thermostatic action of nuclear burning.
  - Contraction of the burning shell would also entail heating and so it must also remain at roughly constant radius.
  - If a core contracts, density and therefore pressure in the shell must decrease to maintain constant temperature and radius.
  - Pressure of the overlying envelope must also decrease.
- Envelope expansion therefore acompanies core contraction.

13.5.1 *Il limite di Schönberg-Chandrasekhar*

Le equazioni dell'equibrio idrostatico e della massa permettono di ottenere:

$$4\pi r^3 \frac{dP}{dM(r)} = -\frac{GM(r)}{r} \tag{13.19}$$

dove il termine a sinistra può essere riscritto nella forma:

$$4\pi r^3 \frac{dP}{dM(r)} = \frac{d\left(4\pi r^3 P\right)}{dM(r)} - 12\pi r^2 P \frac{dr}{dM(r)} = \frac{d\left(4\pi r^3 P\right)}{dM(r)} - \frac{3P}{\rho} \quad .$$

Risostituendo nella (13.19) e integrando sul nucleo isotermo M_{is} si ottiene:

$$\int_0^{M_{is}} \frac{d\left(4\pi r^3 P\right)}{dM(r)} dM(r) - \int_0^{M_{is}} \frac{3P}{\rho} dM(r) = -\int_0^{M_{is}} \frac{GM(r)}{r} dM(r)$$

e quindi (con $M = 0$ per $r = 0$):

$$4\pi R_{is}^3 P_{is} - 3\frac{M_{is}}{\mu_{is} H} kT_{is} = -\frac{3}{5}\frac{GM_{is}^2}{R_{is}}$$

e

$$P_{is} = \frac{3}{4\pi R_{is}^3}\left(\frac{M_{is} kT_{is}}{\mu_{is} H} - \frac{1}{5}\frac{GM_{is}^2}{R_{is}}\right)$$

dove R_{is} e P_{is} sono il raggio del nucleo isotermo e la pressione a tale raggio. Quindi la pressione dipende dai valori specifici di T_{is} e R_{is}. Quando il nucleo isotermo cresce, cresce anche la sua temperatura; il valore massimo della

pressione P_{is} al variare di M_{is} si ha per

$$R_{is} = \frac{2}{5}\frac{GM_{is}\mu_{is}H}{kT_{is}} \tag{13.20}$$

con pressione

$$P_{is,\max} = \frac{365}{64\pi}\frac{1}{G^3 M_{is}^2}\left(\frac{kT_{is}}{\mu_{is}H}\right)^4 \tag{13.21}$$

che indica che la pressione decresce al crescere della massa del nucleo M_{is}. Quindi esiste un valore massimo della massa di un nucleo isotermo per sorreggere la massa sovrastante.

Dobbiamo quindi valutare la pressione dell'inviluppo integrando l'equilibrio idrostatico (assumendo la pressione superficiale nulla):

$$\begin{aligned} P_{is,inv} &= -\int_{P_{is,inv}}^{0} dP = \int_{M_{is}}^{M} \frac{GM(r)}{4\pi r^4} dM(r) \\ &\simeq \frac{G}{8\pi \langle r^4 \rangle}\left(M - M_{is}\right)^2 \end{aligned}$$

dove $\langle r^4 \rangle \simeq R^4/2$ è un valor medio calcolato a qualche valore del raggio della stella. Assumiamo inoltre $M \gg M_{is}$ e otteniamo:

$$P_{is,inv} \simeq \frac{G}{4\pi}\frac{M^2}{R^4} \quad . \tag{13.22}$$

Si utilizza ora l'equazione di stato dei gas perfetti e la condizione che la densità al bordo del nucleo isotermo sia pari alla densità media della stella:

$$\begin{aligned} T_{is} &= \frac{\mu_{inv} H P_{is,inv}}{k \rho_{is,inv}} \\ \rho_{is,inv} &\simeq \frac{M}{4\pi R^3/3} \end{aligned}$$

e si ricava

$$P_{is,inv} \simeq \frac{81}{4\pi} \frac{1}{G^3 M^2} \left(\frac{k T_i}{\mu_{inv} H} \right)^4 \quad . \tag{13.23}$$

Imponendo l'eguaglianza tra le (13.21) e (13.23) si ottiene il limite richiesto:

$$\frac{M_{is}}{M} \approx 0.54 \left(\frac{\mu_{inv}}{\mu_{is}} \right)^2 \tag{13.24}$$

(il calcolo non approssimato cambia il fattore da 0.54 a 0.37).

Stars of 5Msun

When the helium core mass reaches the Schönberg–Chandrasekhar limit, the core collapses.

The collapse restores a non-zero radiative temperature gradient.

The increasing temperature makes hydrogen-shell burning more efficient.

The stellar envelope expands violently and cools down, causing the effective temperature (Teff) to decrease. As a consequence, the opacity increases, leading to convective instability.

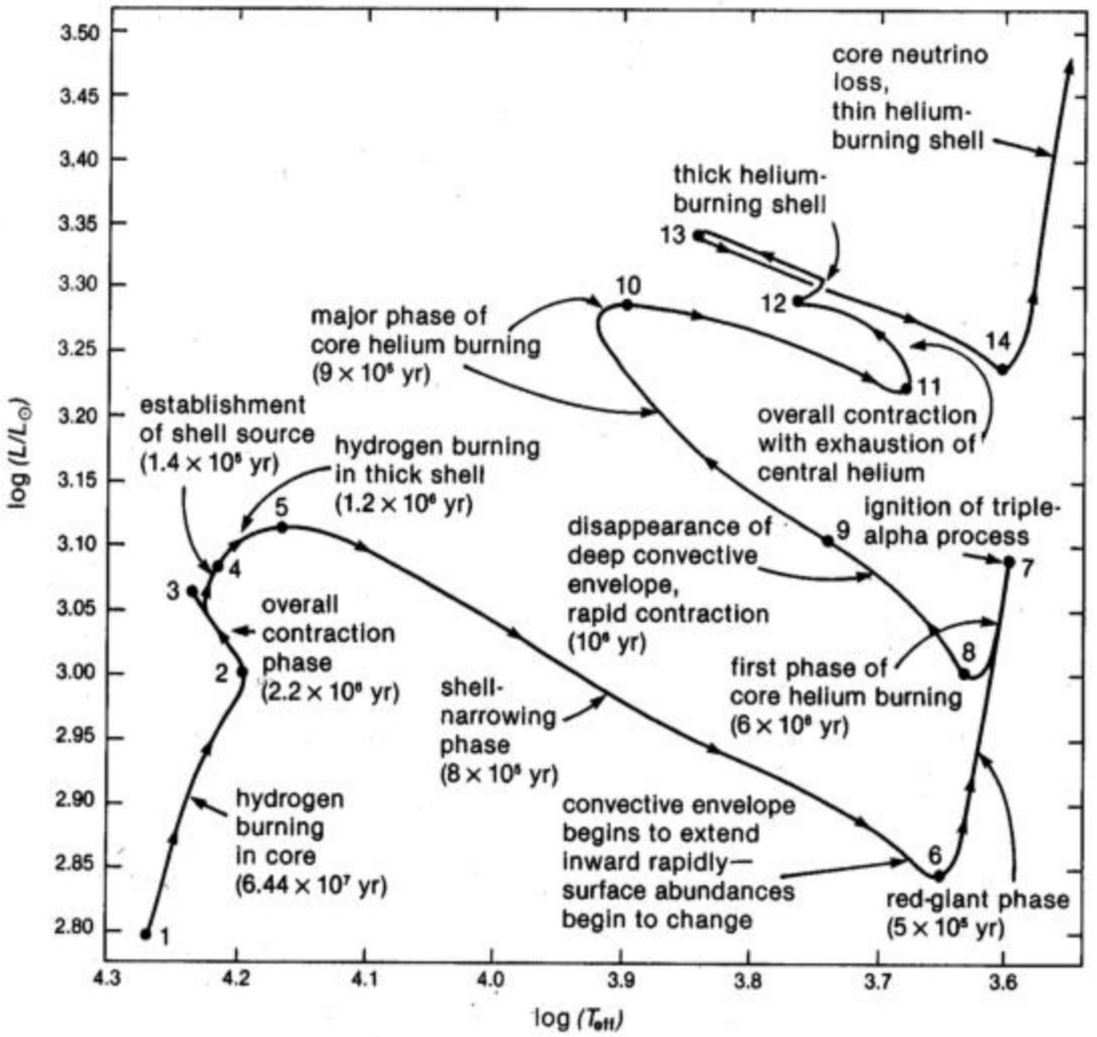
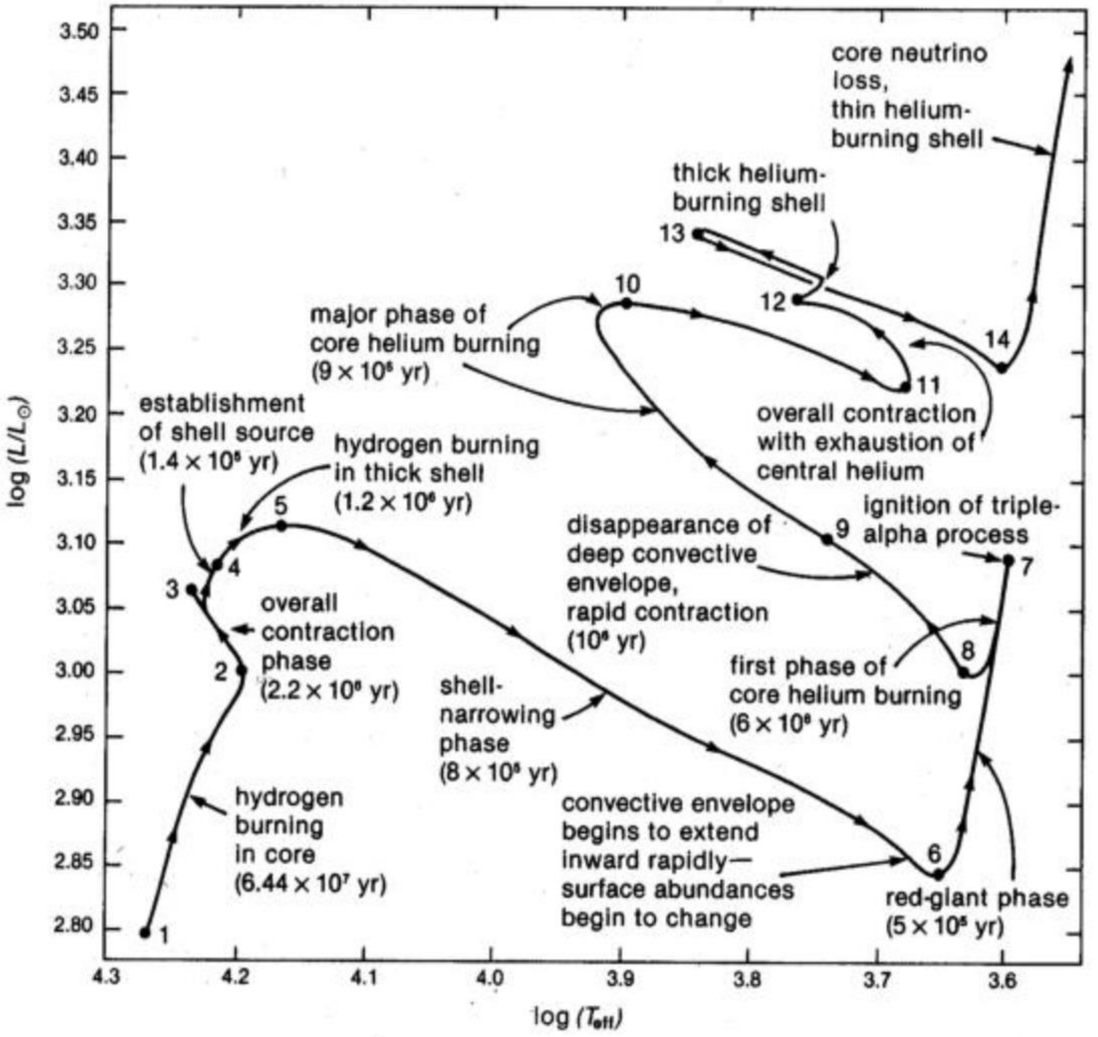


Figure 8.1. Evolutionary path of a Population I star of 5 $M_\odot$ in the Hertzsprung-Russell diagram, showing processes characterizing each stage. Time in parentheses is the duration of the stage between the numbered points. Luminosity is in $L_\odot$ and temperature in K. The elapsed time between points 10 and 11 is 10^6 years.

Stars of 5Msun

This is the so-called dredge-up phase, during which processed material from the stellar interior can be brought up to the surface. This is an important observable effect for validating stellar evolution models.

The work required to expand the envelope also reduces the luminosity that can reach the surface.



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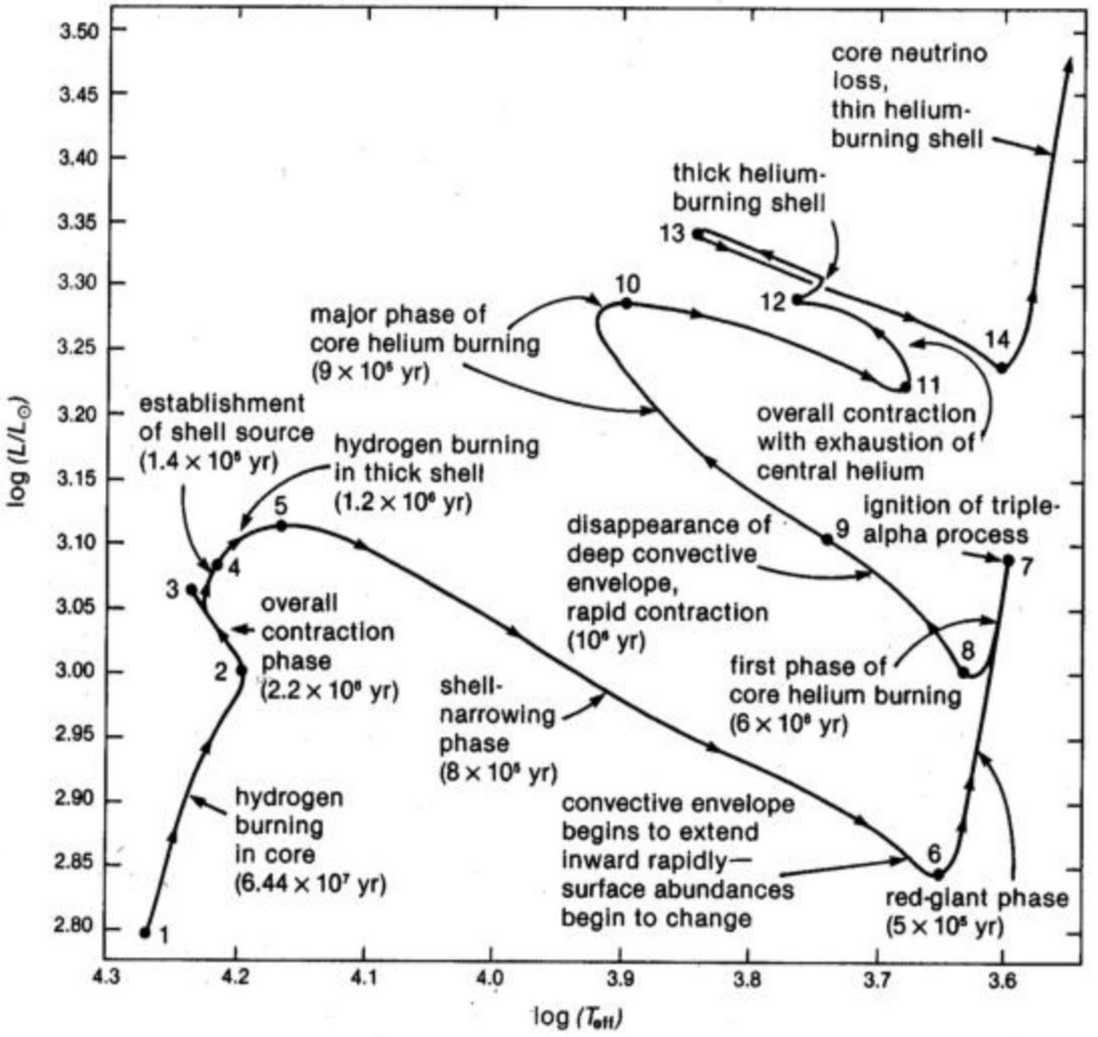
Stars of 5Msun

The star has reached the Hayashi track and therefore evolves along it, increasing the radius with constant Teff. Meanwhile, its temperature in the core is increasing as a result of enhanced hydrogen-shell burning, due to the contraction and increasing density of the shell.

This corresponds to the **red giant branch (RGB) phase.**



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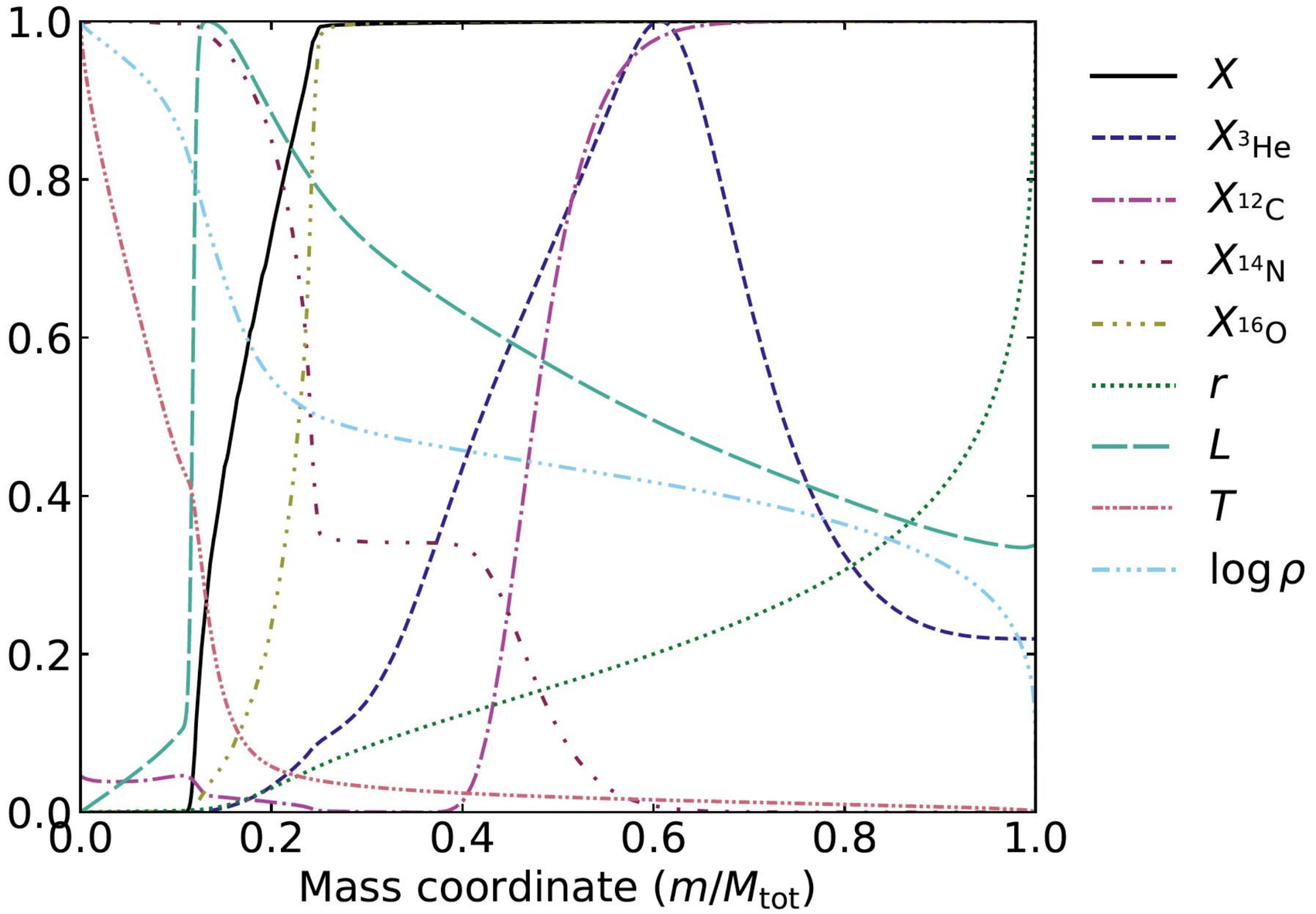
Stars of 5Msun

The variation with mass fraction of a number of quantities in a 5M model just before the first ascent on the Hayashi track



Stars of 5Msun

At a temperature of about 2×10^8 K and a density of roughly 10^3 g cm^{-3}, the conditions for the ignition of helium burning are reached primarily through the reaction

$$3\ He^4 \rightarrow C^{12},$$

and, to a lesser extent,

$$C^{12} + He^4 \rightarrow O^{16}.$$

The core expands, and as a consequence the efficiency of hydrogen-shell burning is partially reduced, although it remains the dominant energy source. The luminosity decreases, but at that point the star contracts and reaches a new equilibrium at a higher temperature.

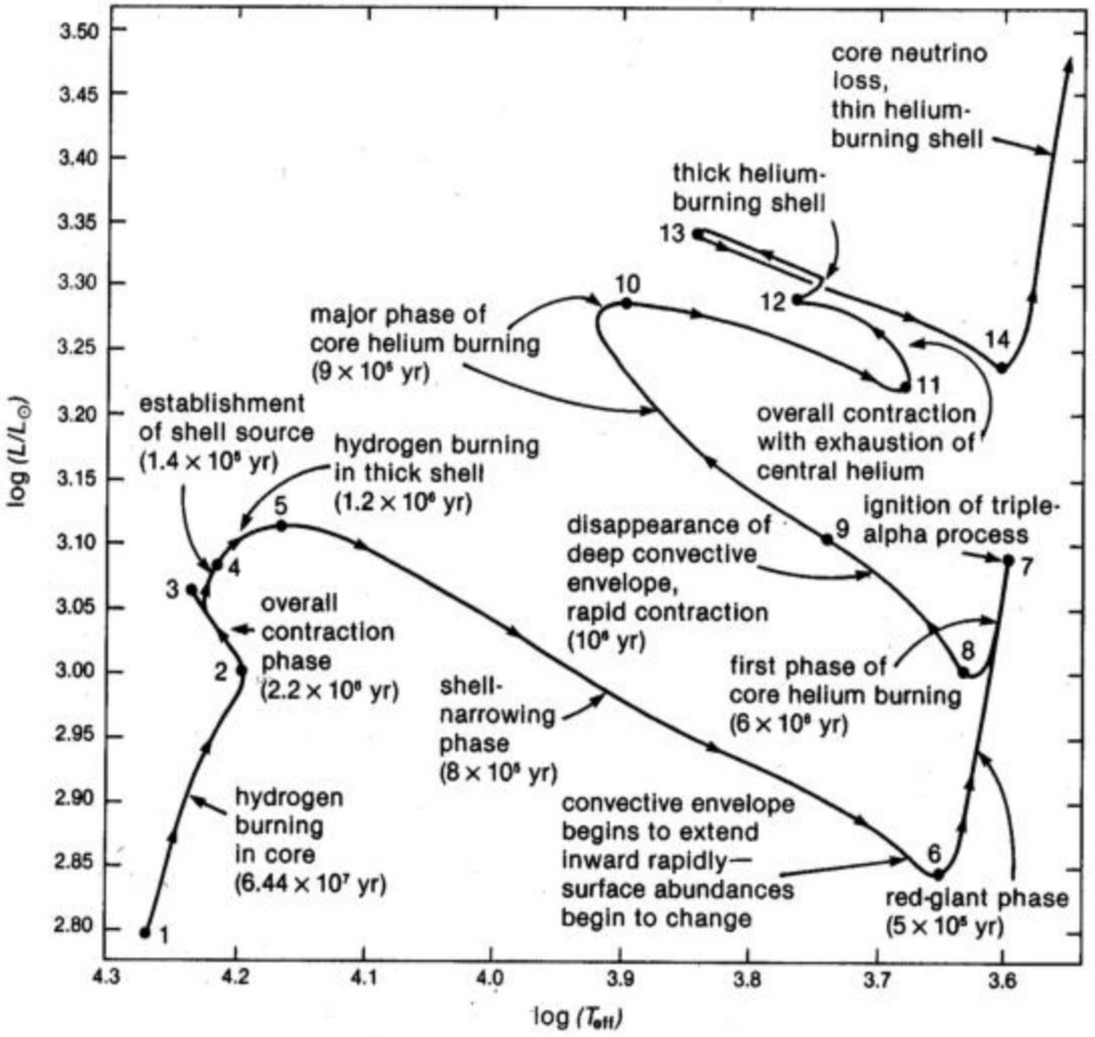
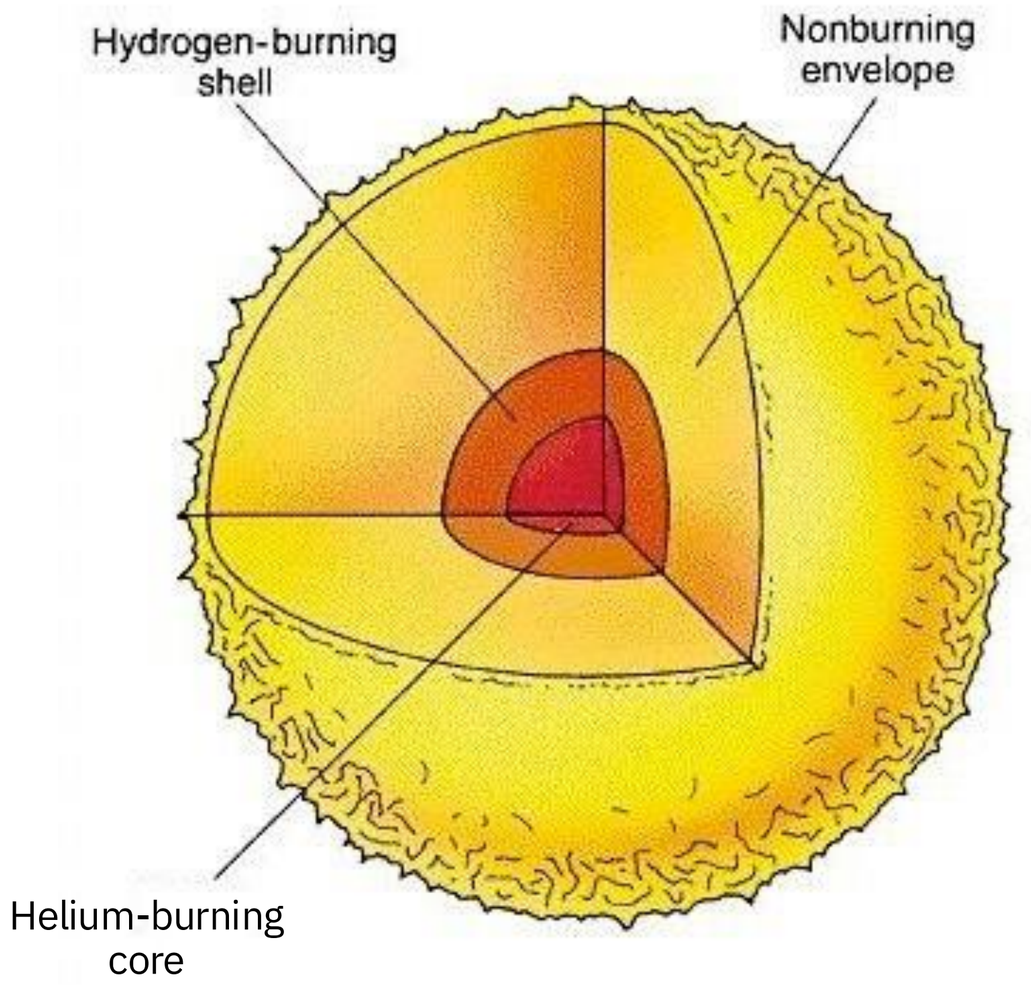


Figure 8.1. Evolutionary path of a Population I star of 5 $M_\odot$ in the Hertzsprung-Russell diagram, showing processes characterizing each stage. Time in parentheses is the duration of the stage between the numbered points. Luminosity is in $L_\odot$ and temperature in K. The elapsed time between points 10 and 11 is 10^6 years.

The duration of the helium-burning phase is much shorter than that of hydrogen burning: a 5-$M_{\odot}$ star exhausts the helium in its core in about 9 million years.



Stars of 5Msun

The region of the HR diagram where helium burns in the core is called the **horizontal branch (HB).**

Along this branch, stars may become dynamically unstable, developing periodic oscillations and undergoing mass loss. Convection forms again in the envelope, producing a **second dredge-up**, which increases the surface helium abundance.

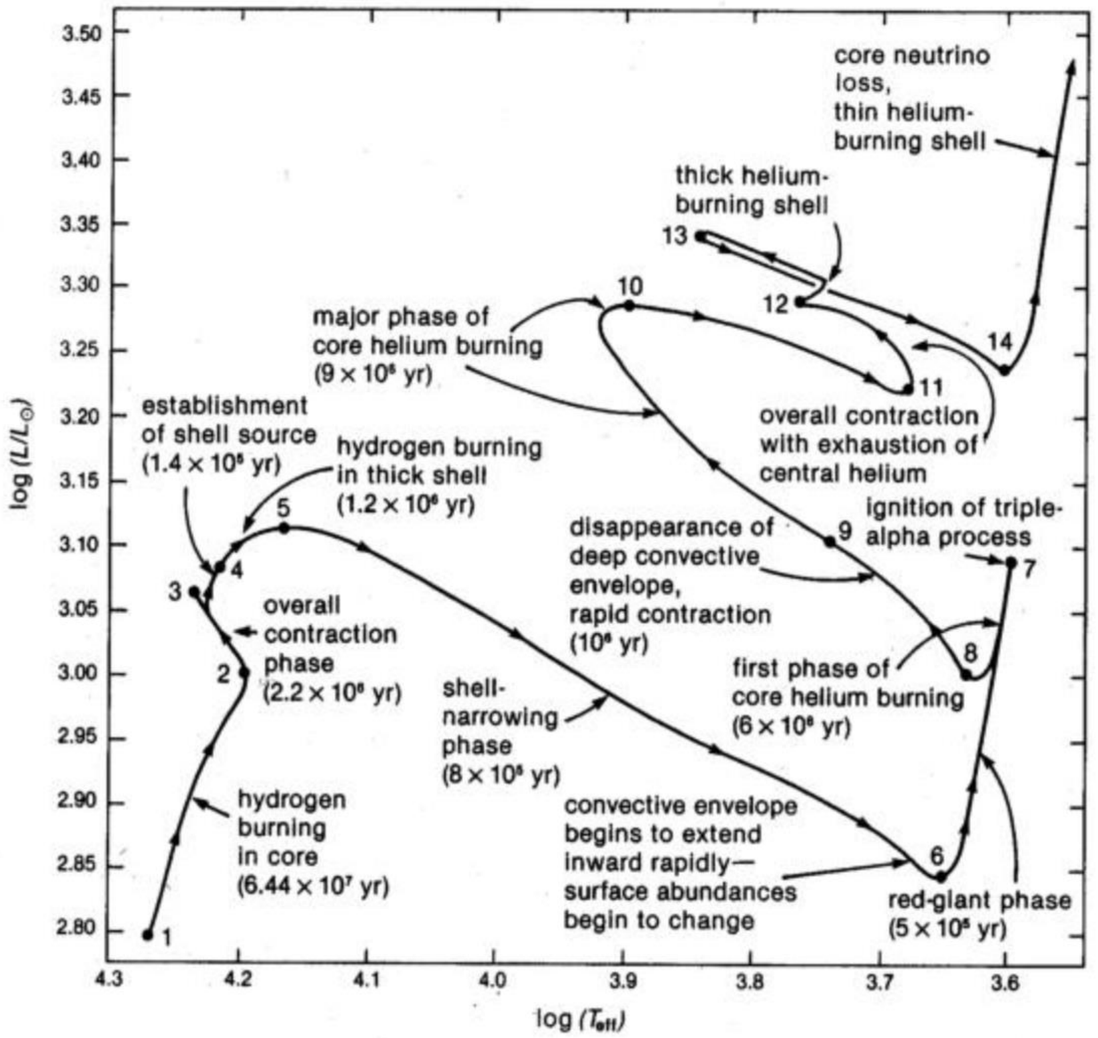


Figure 8.1. Evolutionary path of a Population I star of 5 $M_\odot$ in the Hertzsprung-Russell diagram, showing processes characterizing each stage. Time in parentheses is the duration of the stage between the numbered points. Luminosity is in $L_\odot$ and temperature in K. The elapsed time between points 10 and 11 is 10^6 years.

Stars of 5Msun

When the helium in the core is exhausted, the CO core undergoes the same contraction phase experienced at the end of hydrogen burning, leading to the formation of a shell in which helium burning continues. During this phase of helium ignition and exhaustion in the core, the star moves away from the Hayashi track toward the high-temperature region of the HR diagram, and then essentially returns to the Hayashi track, although at a higher temperature.

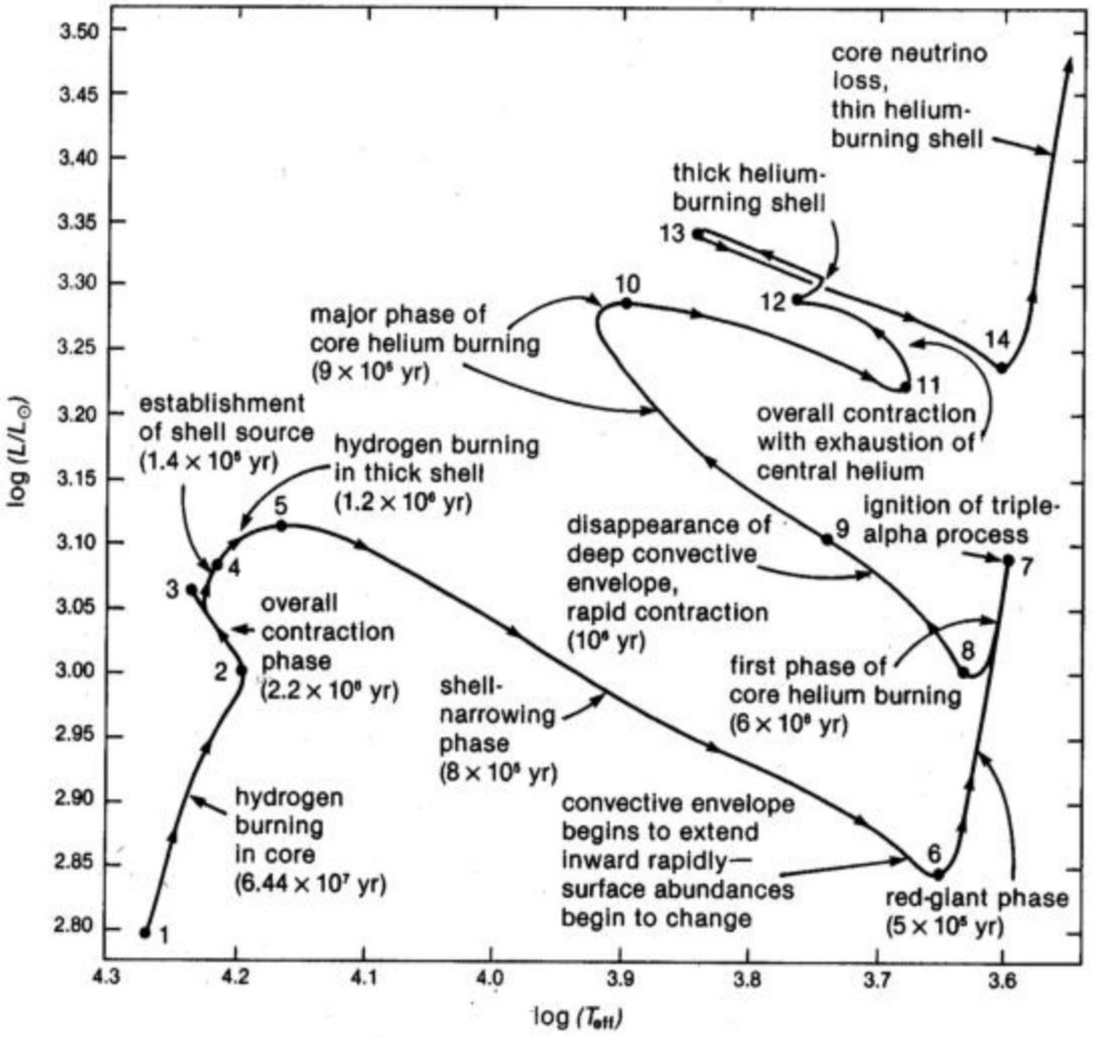
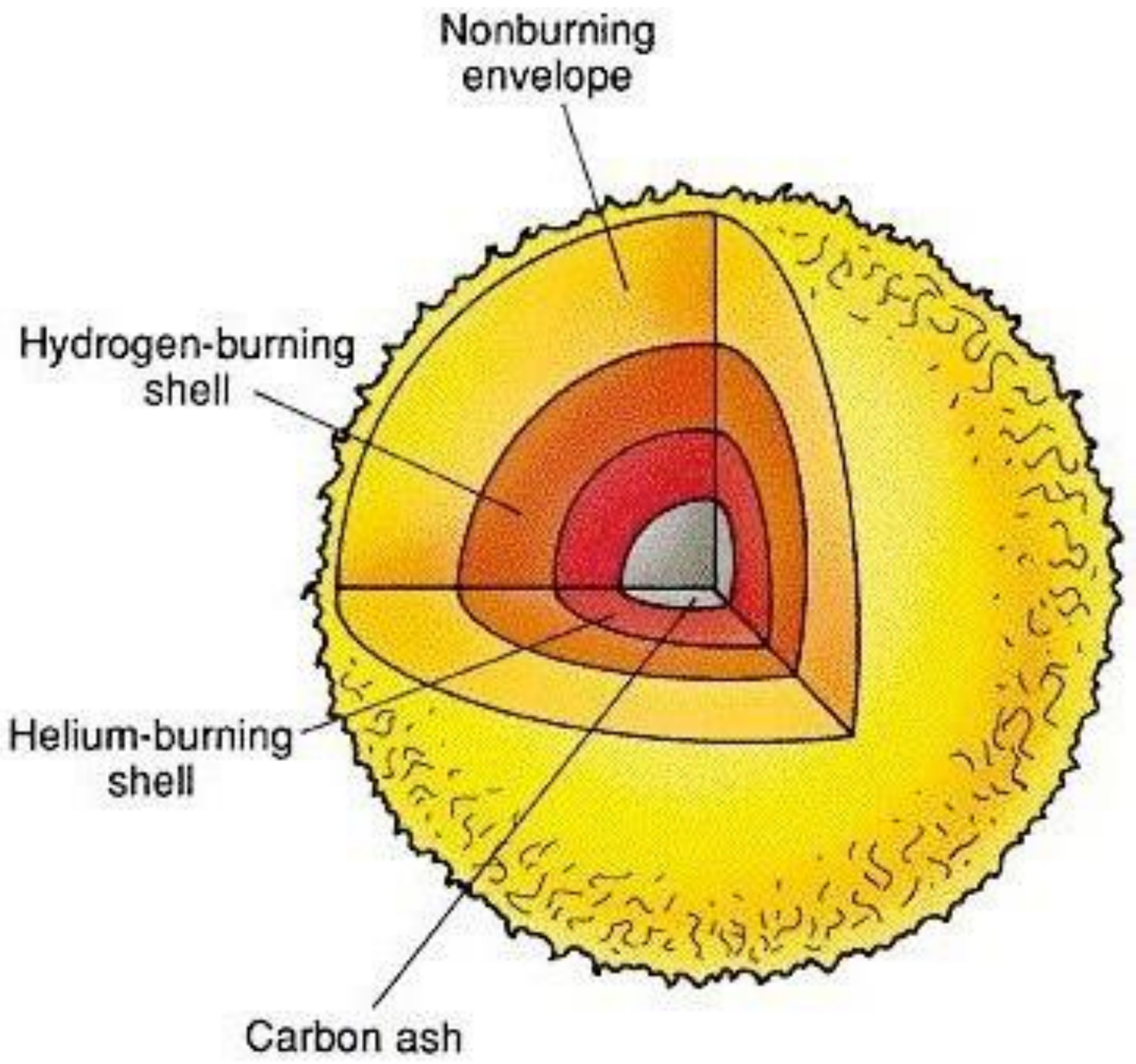


Figure 8.1. Evolutionary path of a Population I star of 5 $M_\odot$ in the Hertzsprung-Russell diagram, showing processes characterizing each stage. Time in parentheses is the duration of the stage between the numbered points. Luminosity is in $L_\odot$ and temperature in K. The elapsed time between points 10 and 11 is 10^6 years.

Stars of 5Msun

When the star, now with a carbon–oxygen core and two active burning shells (helium and hydrogen), reaches the Hayashi track, its evolutionary track moves upward along the **asymptotic giant branch (AGB).**

The core temperature is still on the order of 2×10^8 K, but the density is close to 10^6 g cm^{-3}.

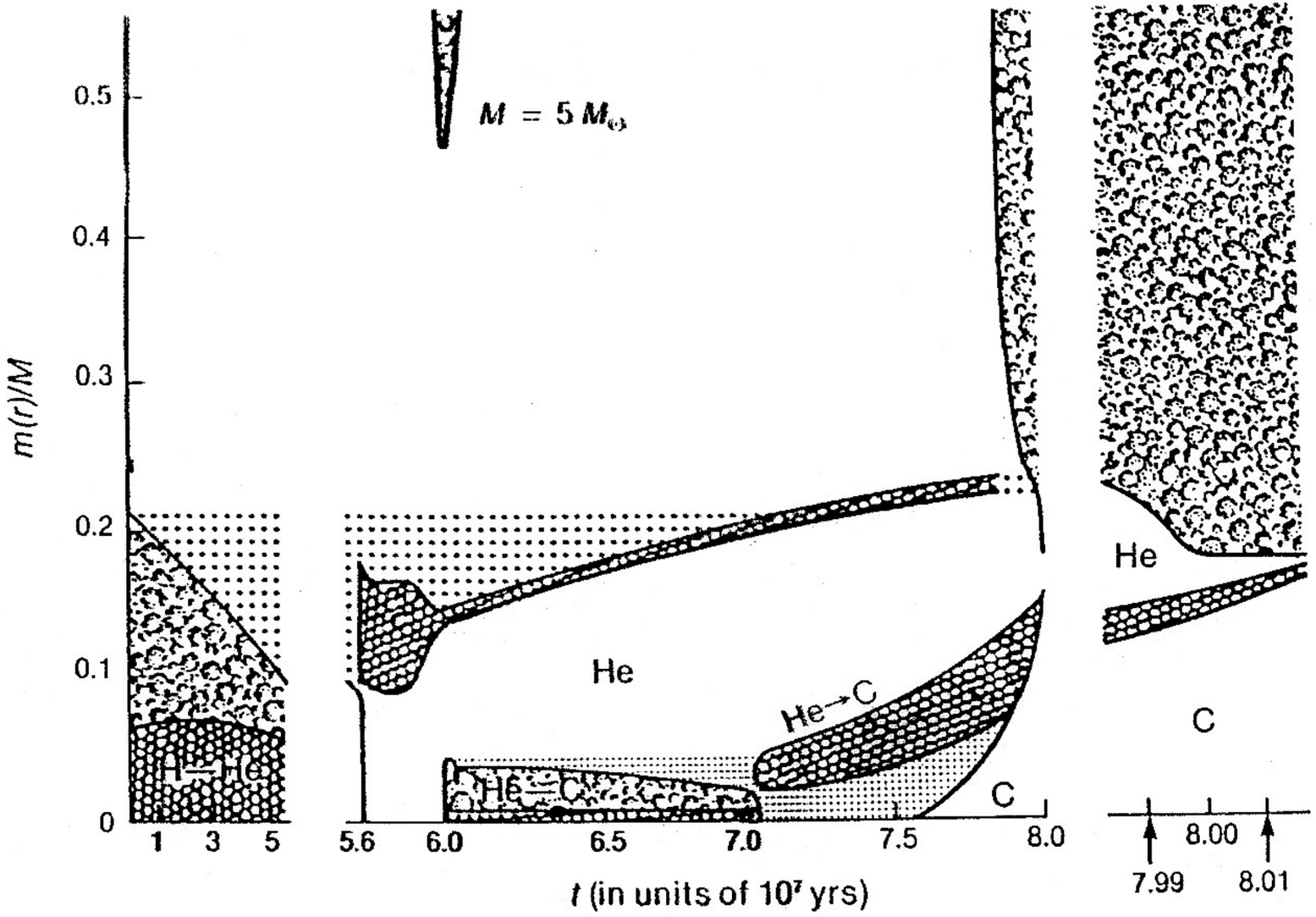


Stars of 5Msun

Internal structure in different phases of evolution

Internal structure of a 5-$M_\odot$ star in the post-main-sequence phase.

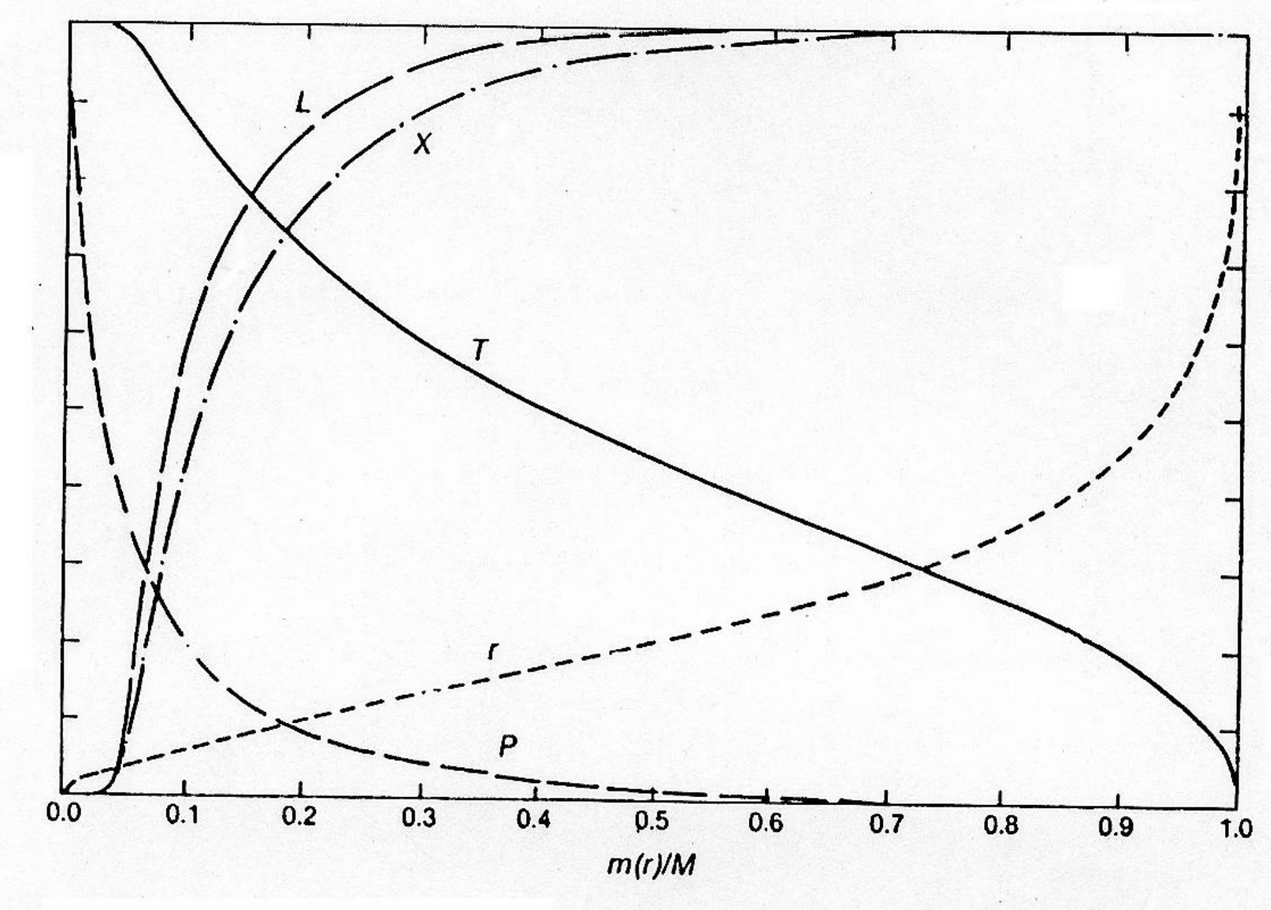
Convective transport regions are indicated by curly patterns, thermonuclear burning regions by honeycomb patterns, and zones with changing chemical composition by dots — sparse for H → He, dense for He → C.



Post main sequence for stars of M<2Msun (Sun!)

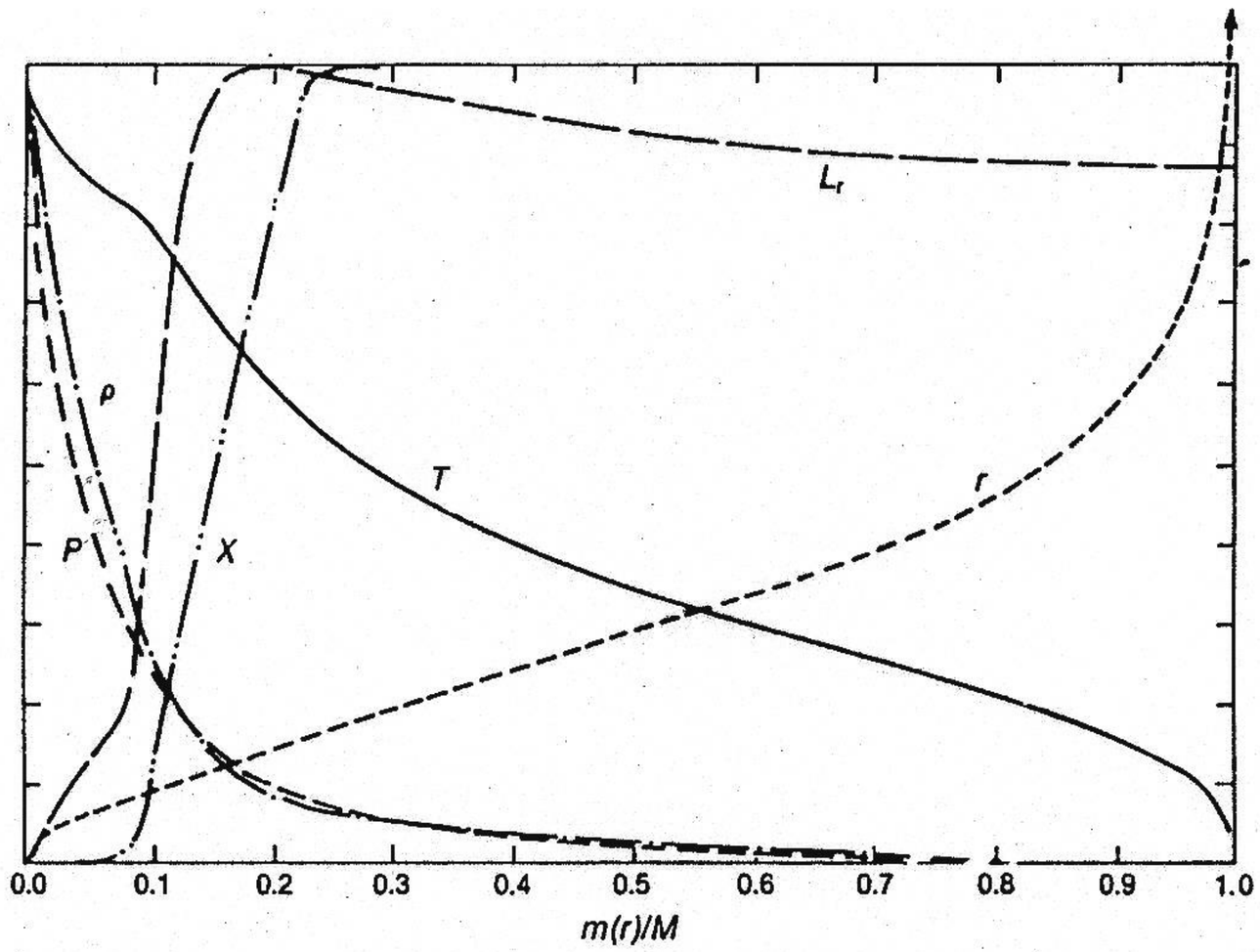
Similar to higher-mass stars.
The differences are that energy transport is radiative, and therefore, no mixing occurs. The core contracts slightly under quasi-hydrostatic conditions to compensate for radiative losses, and thermonuclear burning ignites in the circum-nuclear shell, producing a luminosity even higher than during core burning.

Here as well, the core becomes nearly isothermal.



elapsed time since the initial model for the phase before the main sequence is 9.2015×10^9 years. The lower limits of the ordinate are zero. The upper limit of the ordinate for each curve is $r = 1.268\ R_\odot$; $P_c = 1.315 \times 10^{18}$ dynes cm^{-2}; $T_c = 1.91 \times 10^7$ K; $L = 2.13\ L_\odot$; $X_s = 0.708$. The actual stellar radius is $R = 1.353\ R_\odot$, and the central density is 1026.0 g cm^{-3}.

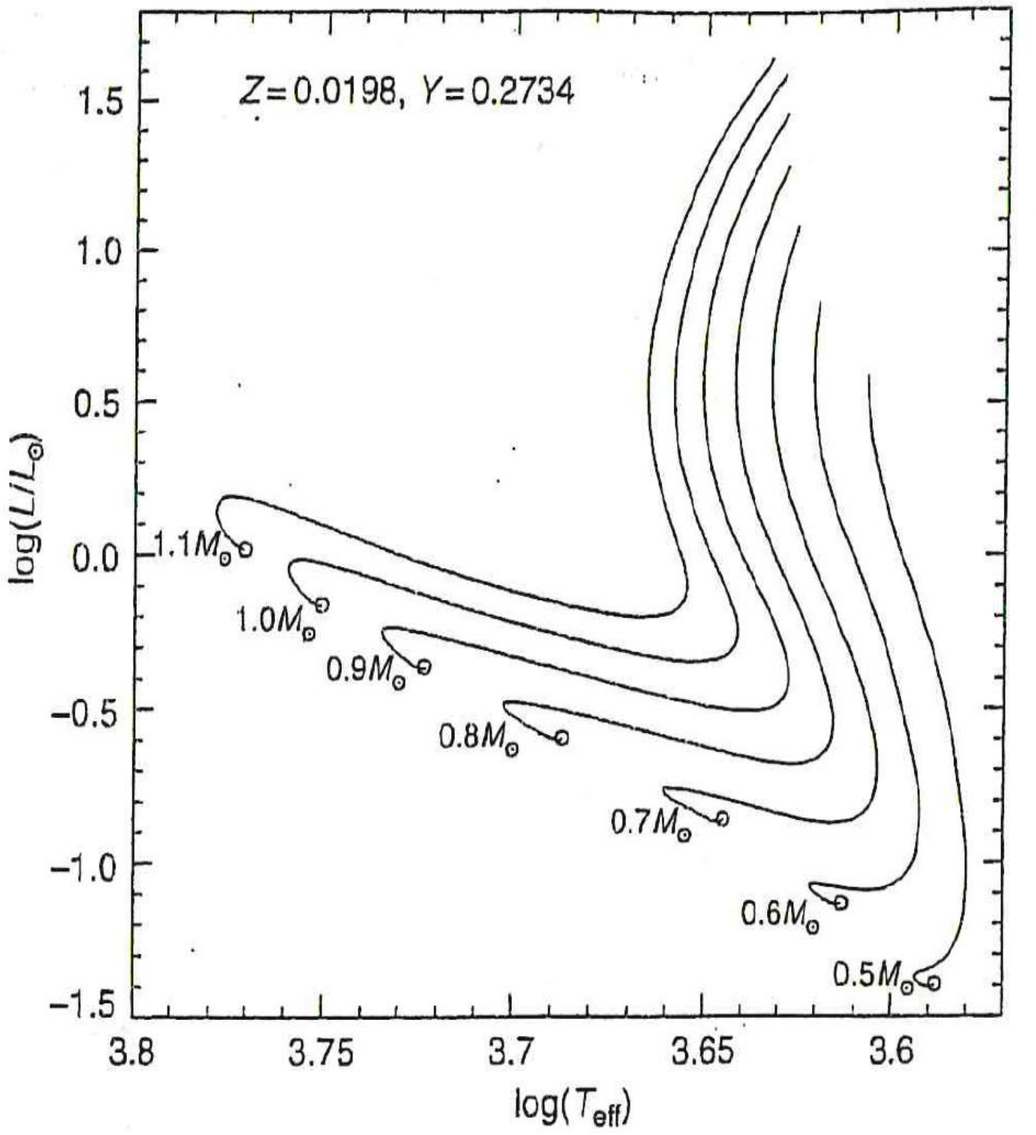
Stars of 5Msun (COMPARISON)



Model of a 5 $M_\odot$ star just after it leaves the main sequence at a time of 6.84461×10^7 years. Maximum value of the ordinate for each curve; $r = 2.9198\ R_\odot$; $P_c = 2.427 \times 10^{17}$ dynes cm^{-2}; $\rho_c = 106.59$ g cm^{-3}; $T_c = 3.6164 \times 10^7$ K; $L = 1.291 \times 10^3\ L_\odot$; and $X_c = 0.708$. The total radius is $R = 3.943\ R_\odot$.

Post main sequence for stars of M<2Msun (Sun!)

Initially, the evolutionary track moves upward and to the left along the main sequence line; it then bends toward lower effective temperatures because the high opacity prevents all the produced luminosity from being transported outward, and part of the energy is instead used to perform work in the slow expansion of the envelope. This portion of the evolutionary track, shifting toward the red in the HR diagram, is referred to as the subgiant branch (SGB).



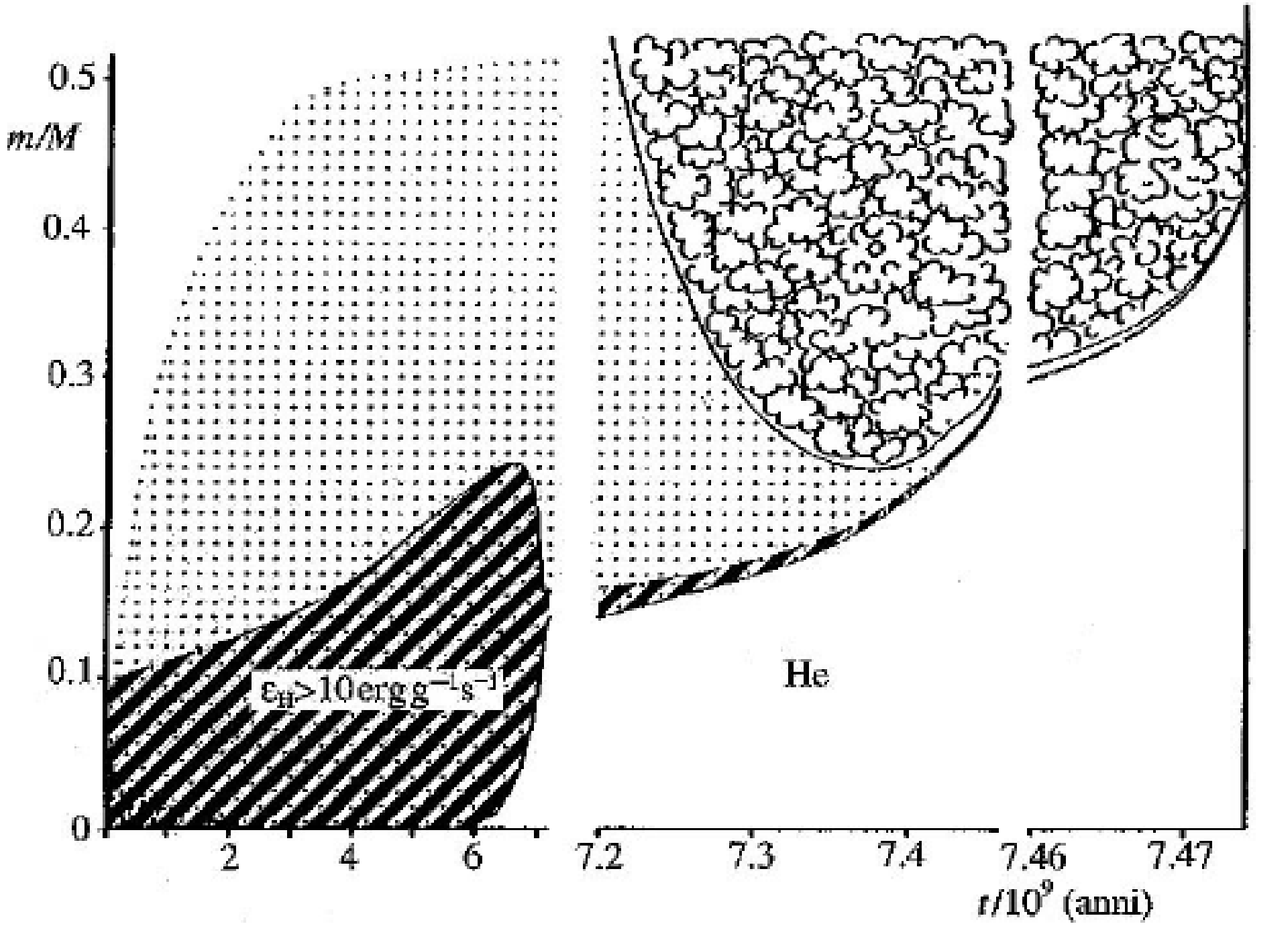
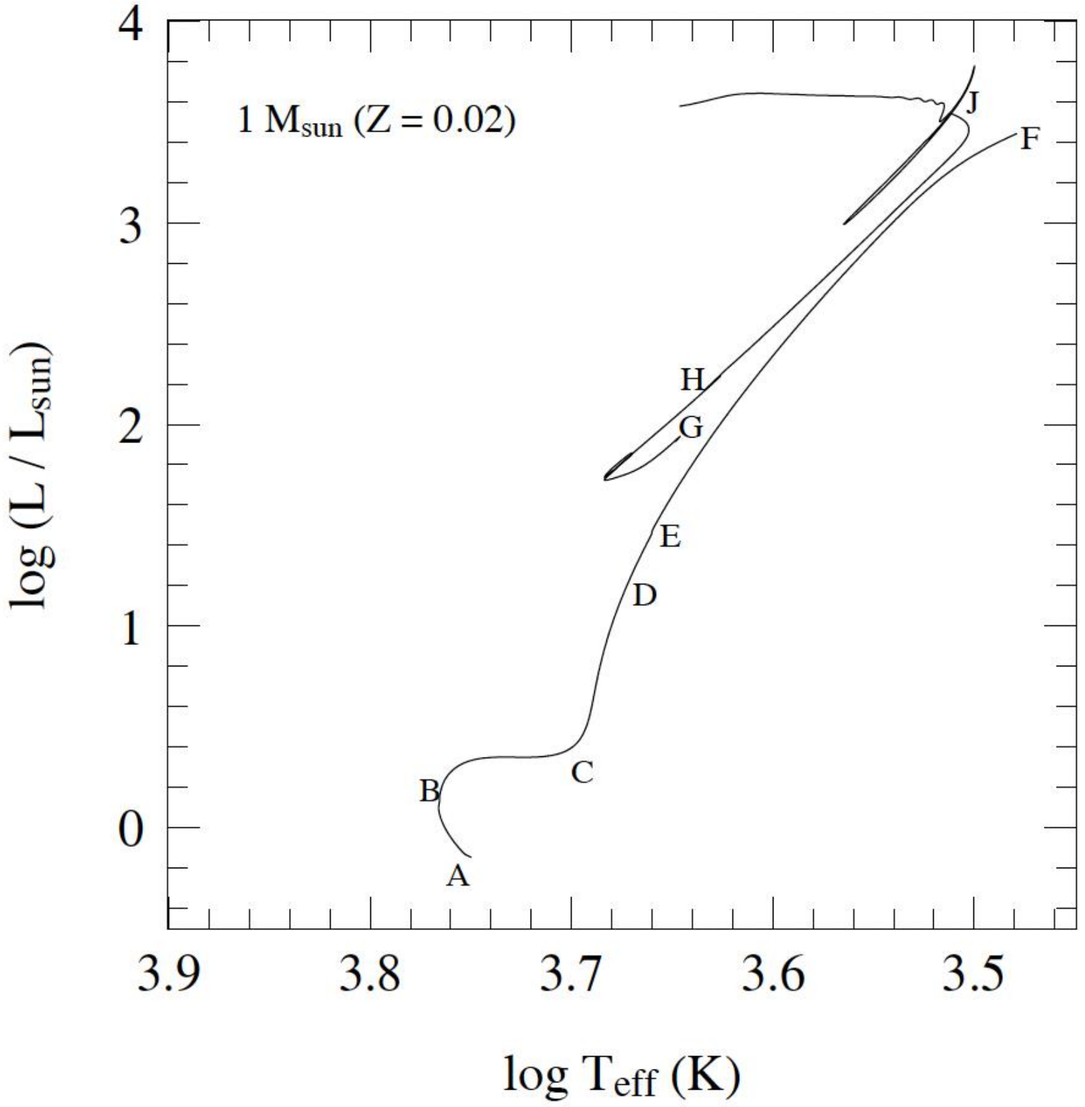


Fig. 13.22: Struttura interna di una stella di $1.3M_\odot$ nella fase post-MS. regioni di trasporto convettivo sono mostrate con ricciolini, regioni con bruciamento termonucleare $H \longrightarrow He$ con tratteggi spessi, regioni bruciamenti parziali con puntini.

Post main sequence for stars of M<2Msun (Sun!)

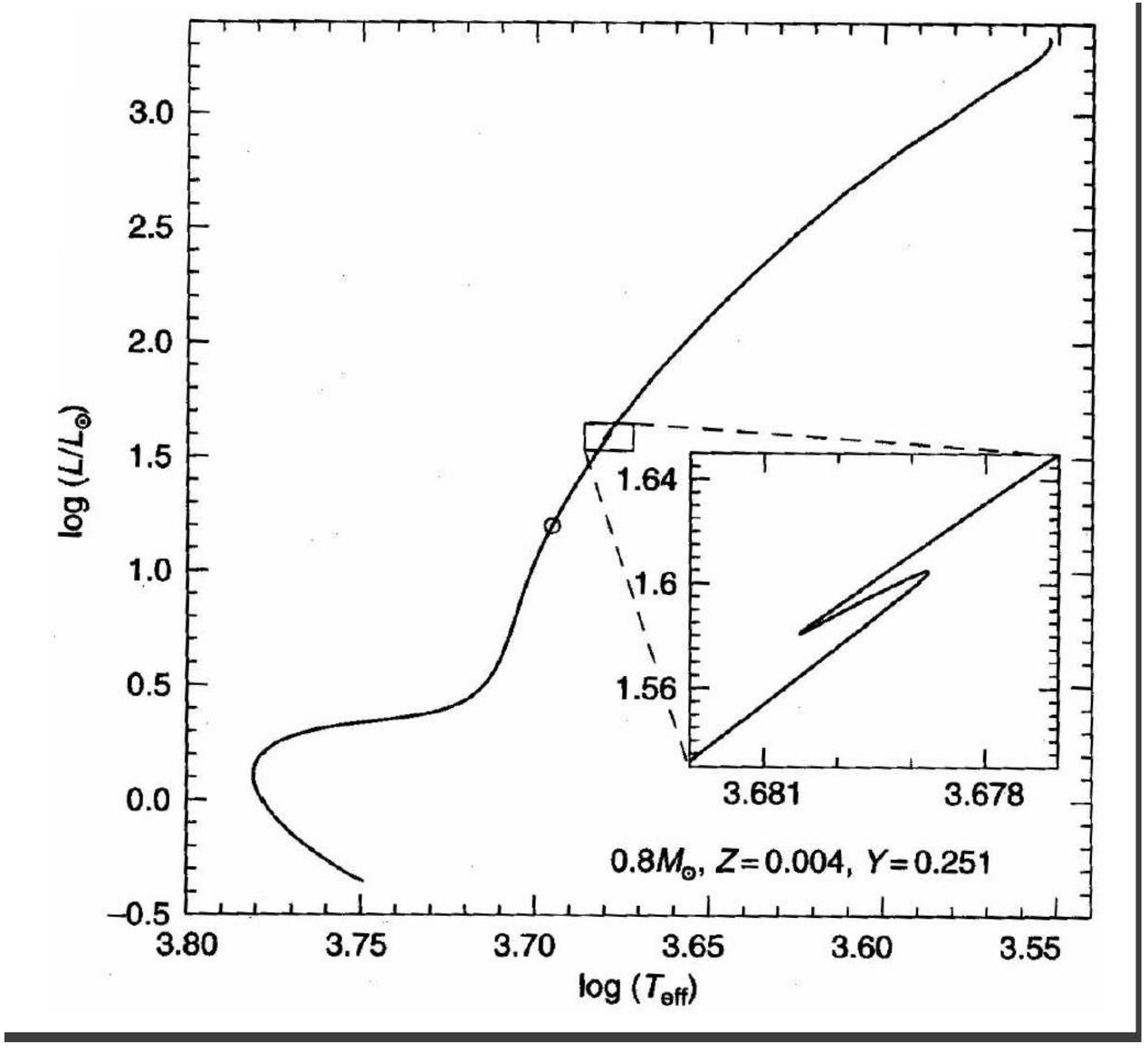
- The star is forced to move almost vertically upward in the HR diagram and becomes a red giant.
- The increase in luminosity, which can no longer be dissipated through expansion alone, triggers convection in the giant's envelope.
- At the same time, the core continues to contract and becomes degenerate.



Post main sequence for stars of M<2Msun (Sun!)

First Dredge-Up and the Luminosity Bump:

- When the convective envelope reaches its deepest extent at D, it has penetrated into layers processed by H-burning during the Main Sequence and have been partly processed by the CN-Cycle.
- Up to D the surface He-abundance increases and the H-abundance decreases. More noticeably the C/N ratio decreases by a large factor; this is called the **First Dredge-Up** phase.
- Later at E, the H-burning shell has burnt its way out to the discontinuity left by the convective envelope at its deepest extent.
- The H-burning shell is then quite suddenly in an environment with a higher H-abundance (and lower mean molecular weight).
- As a result, the H-burning shell burns at a slightly lower rate leading temporarily to a slightly lower luminosity as shown in the next slide.
- The resulting loop (the star crosses this luminosity range three times) results in a larger number of stars in this luminosity range in a stellar population.

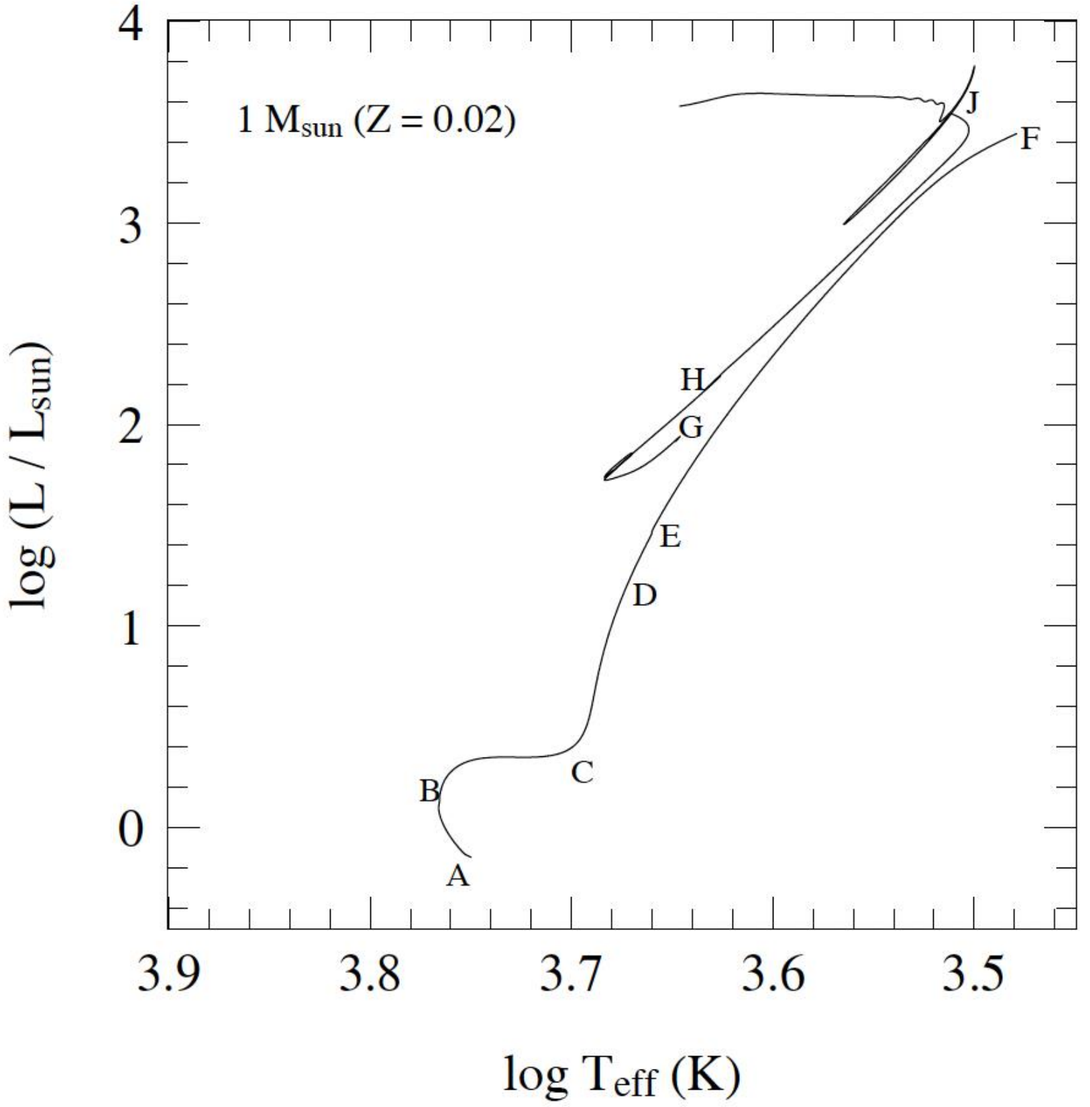

3.0
2.5
2.0
1.5
1.0
0.5
0.0
−0.5
log ($L/L_\odot$)
3.80
3.75
3.70
3.65
3.60
3.55
log (T_{eff})
1.64
1.6
1.56
3.681
3.678
$0.8M_\odot$, $Z=0.004$, $Y=0.251$

Post main sequence for stars of M<2Msun (Sun!)

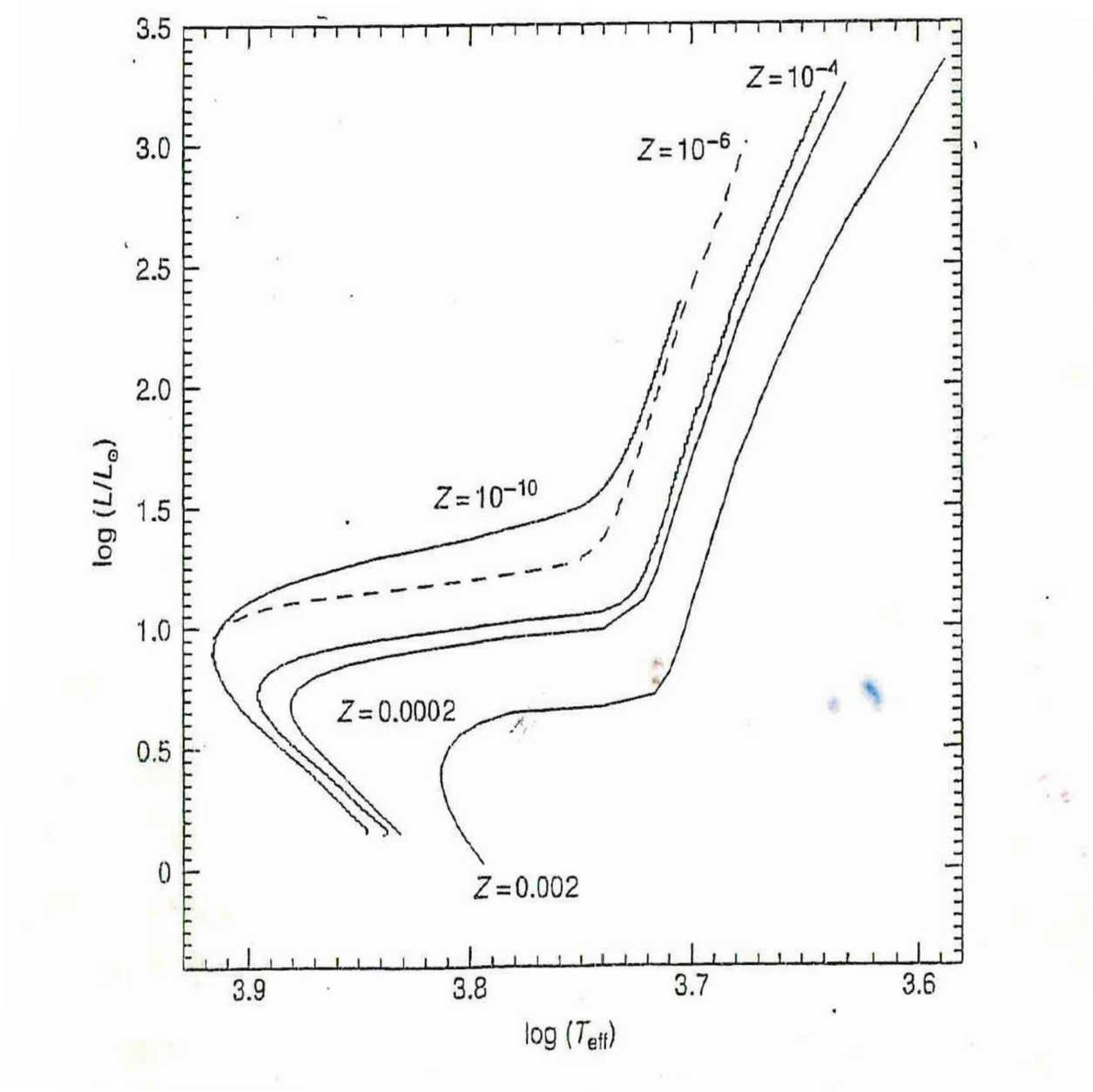
Helium ignition in the core occurs explosively (**core helium flash**) at a temperature of about $\mathbf{2 \times 10^8\ K}$ (as in the case of a 5-M⊙ star) and at a density of approximately $\mathbf{10^6\ g\ cm^{-3}}$.

The ignition is explosive because the core is degenerate: as the temperature increases, the pressure does **not** increase accordingly.

The temperature continues to rise until the material is no longer degenerate, at which point a violent expansion and explosive event follow.



Dependence to metallicity for stars of 1Msun

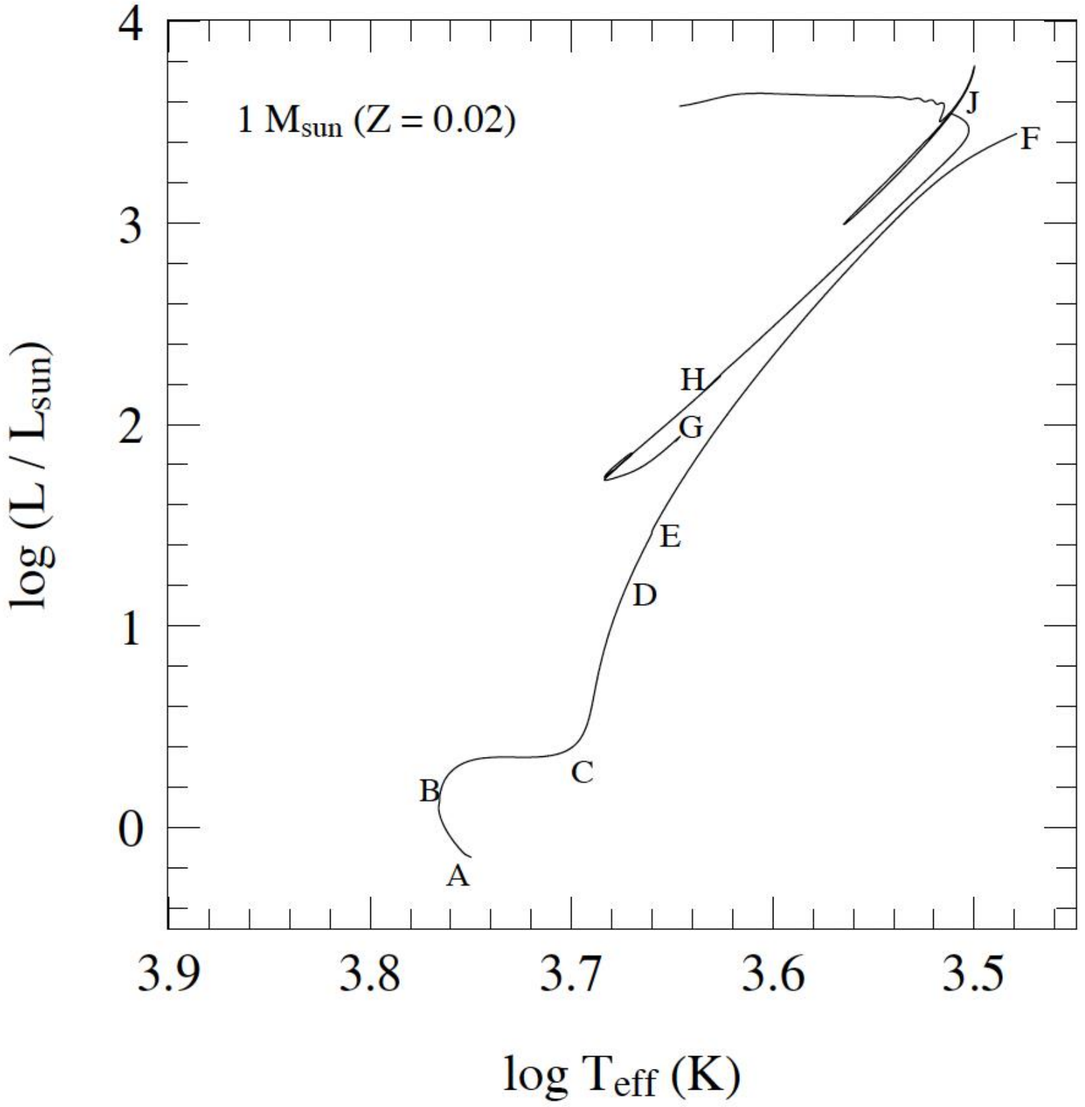


Post main sequence for stars of M<2Msun (Sun!)

- The star burns helium in the core and hydrogen in a shell. The shell burning is less efficient than before, due to the expansion caused by the helium flash.
  Helium burning begins when the helium core mass reaches 0.48–0.50 $M_{\odot}$.

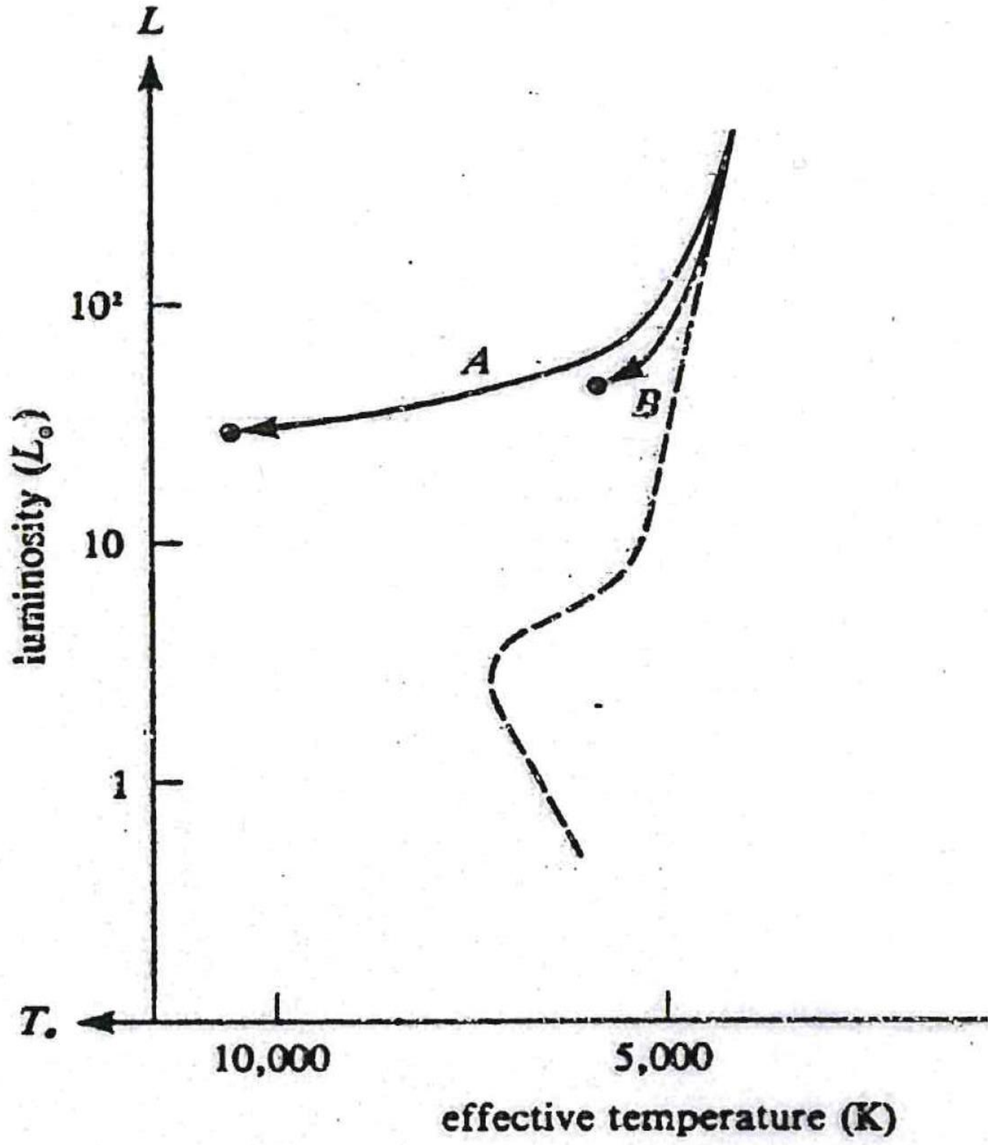
- The intensity of hydrogen-shell burning depends on the mass overlying the shell.

- This phase of double burning is referred to as the horizontal branch (HB).



Post main sequence for stars of M<2Msun (Sun!)

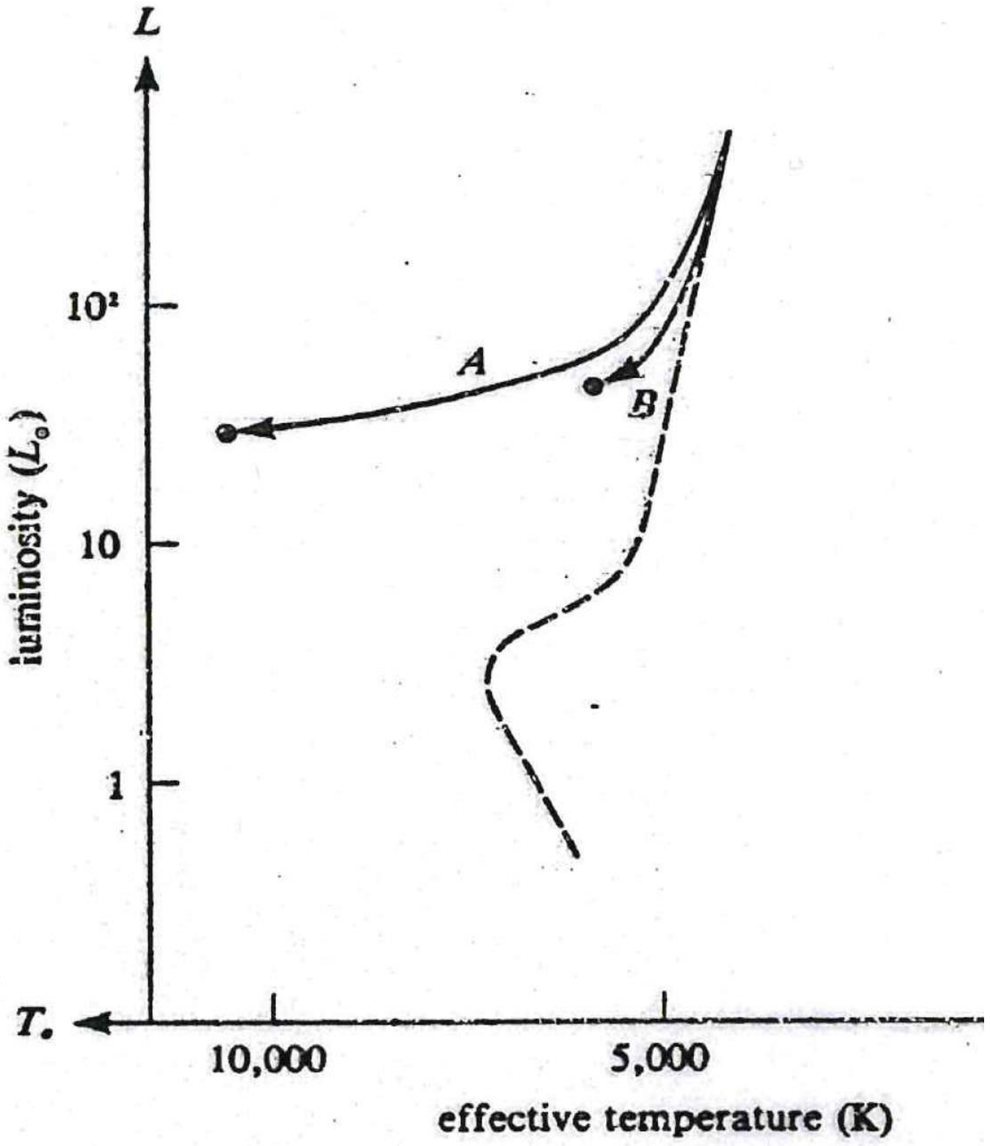
A low-metallicity (Population II) star moves down along the HB depending on the mass loss it experienced during the helium flash. The greater the mass loss, the smaller the envelope, and the higher the effective temperature (case A).



Post main sequence for stars of M<2Msun (Sun!)

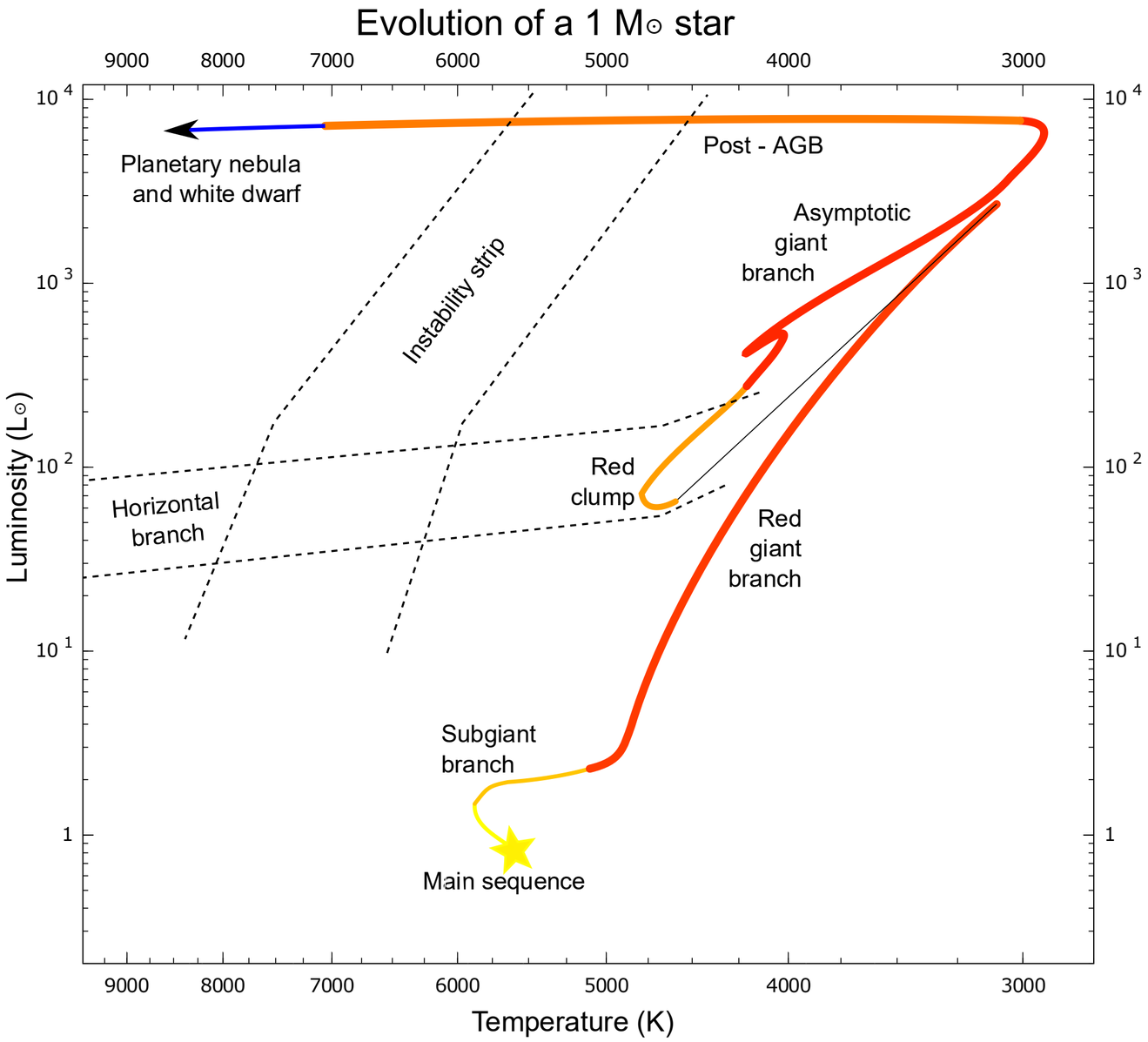
If the metal mass fraction Z is low, the HB becomes more extended (case A). In fact, it is observed only in globular clusters (low Z).

The HB phase lasts for about 100 million years, and these stars can provide a significant contribution to the integrated blue light of elliptical galaxies.



Red clump

High-metallicity (Population I) stars move down the HB at very low effective temperatures and may therefore be confused with stars on the red-giant (RG) phase. They are recognizable because they form a distinct clump, known as the red clump.



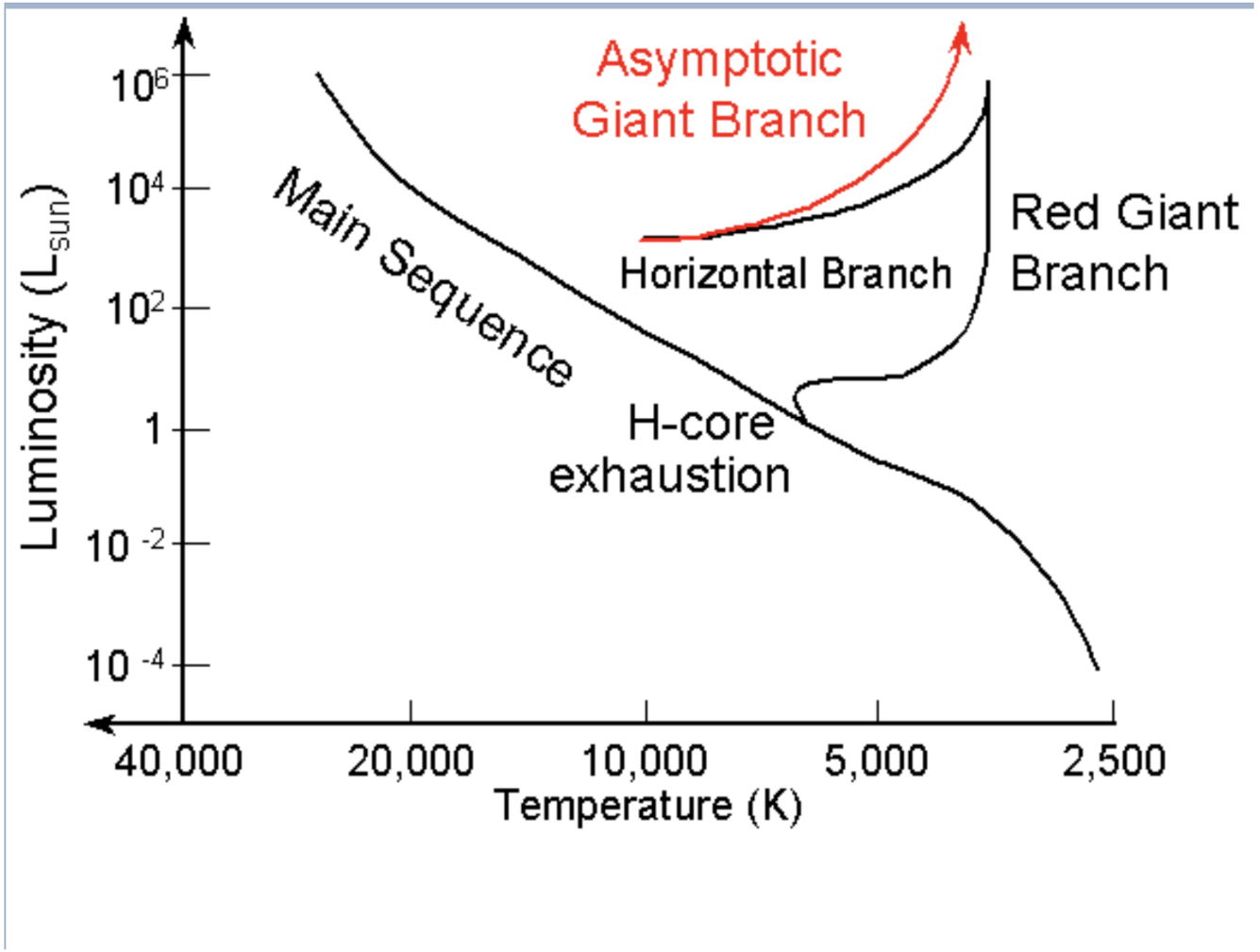
Ascent to the asymptotic giant branch (AGB)

(similar for $0.5 < M < 6\ M_{\odot}$)

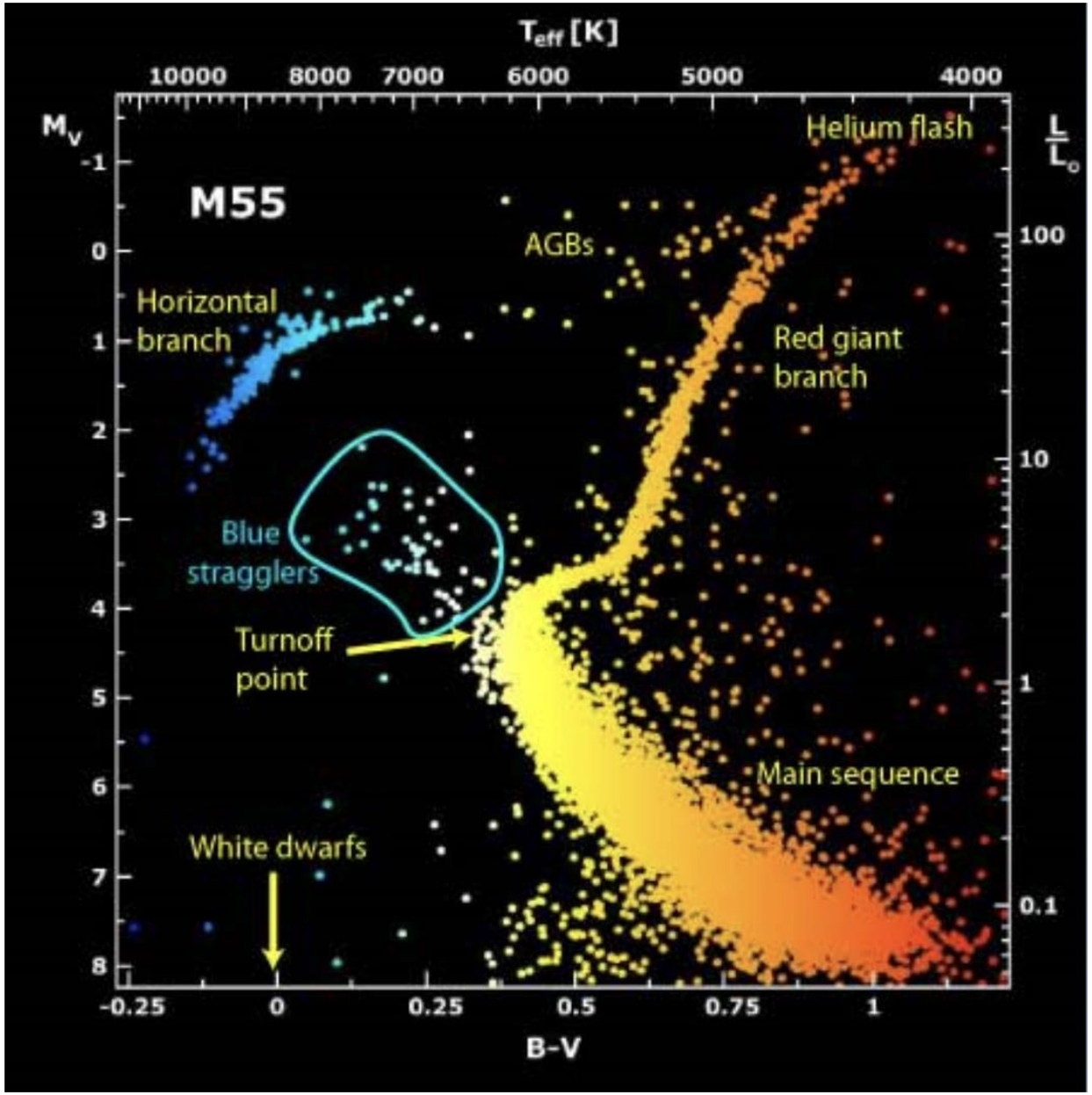
When helium in the core is exhausted, the carbon–oxygen core begins to contract, and helium ignites in a shell, as already seen for more massive stars.

The star now burns hydrogen in one shell and helium in another shell.

This corresponds to the double-shell burning phase.

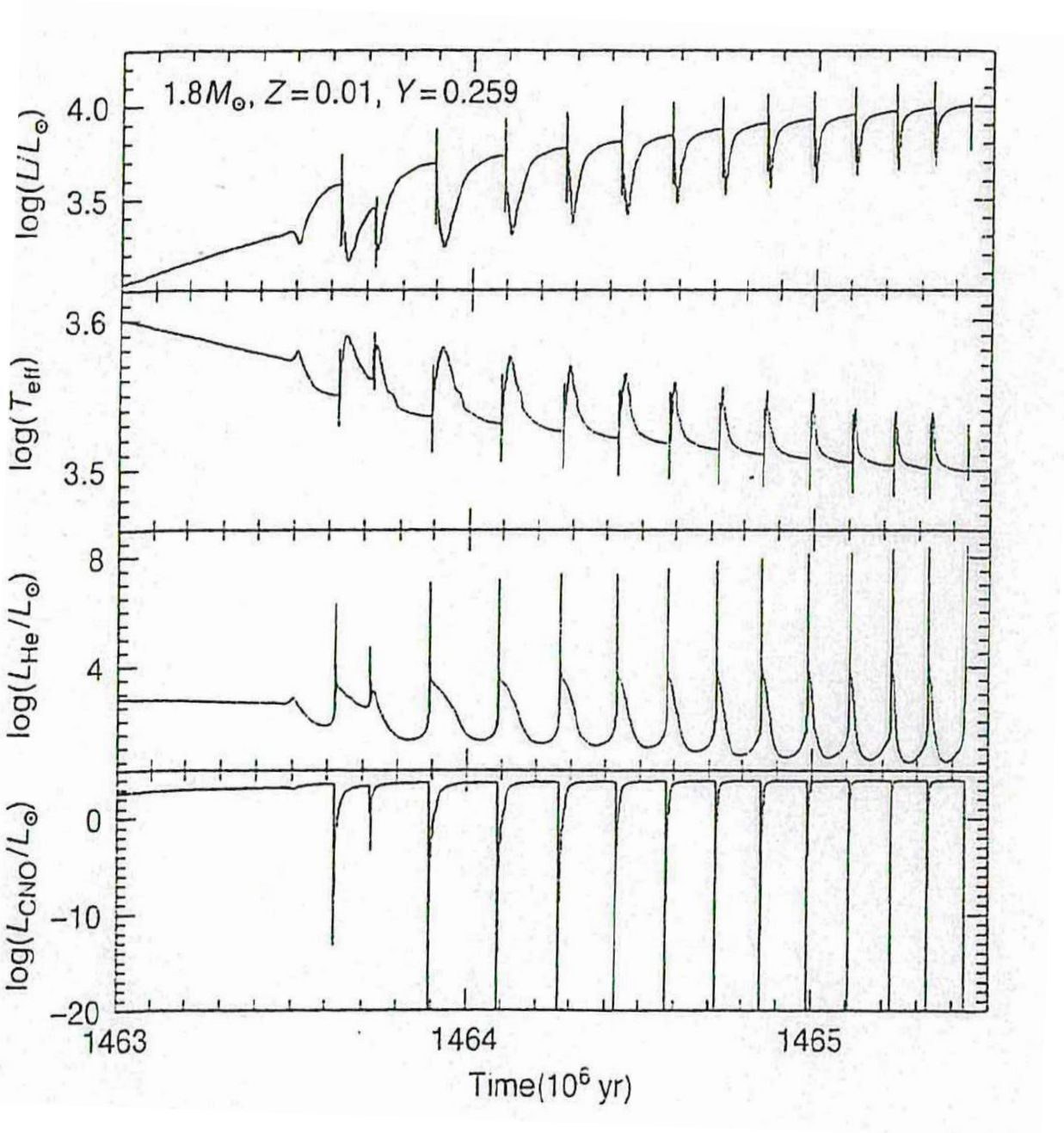


Blue stragglers?



AGB phase

The activity of the helium shell is intermittent because it is continuously replenished by the overlying hydrogen-burning shell. This produces the so-called **helium shell flash,** caused by the fact that the gas is close to the limit of electron degeneracy.
To re-ignite the helium shell each time, the degeneracy must first be lifted through contraction, while the increase in luminosity produced by the flash causes the shell to expand again.

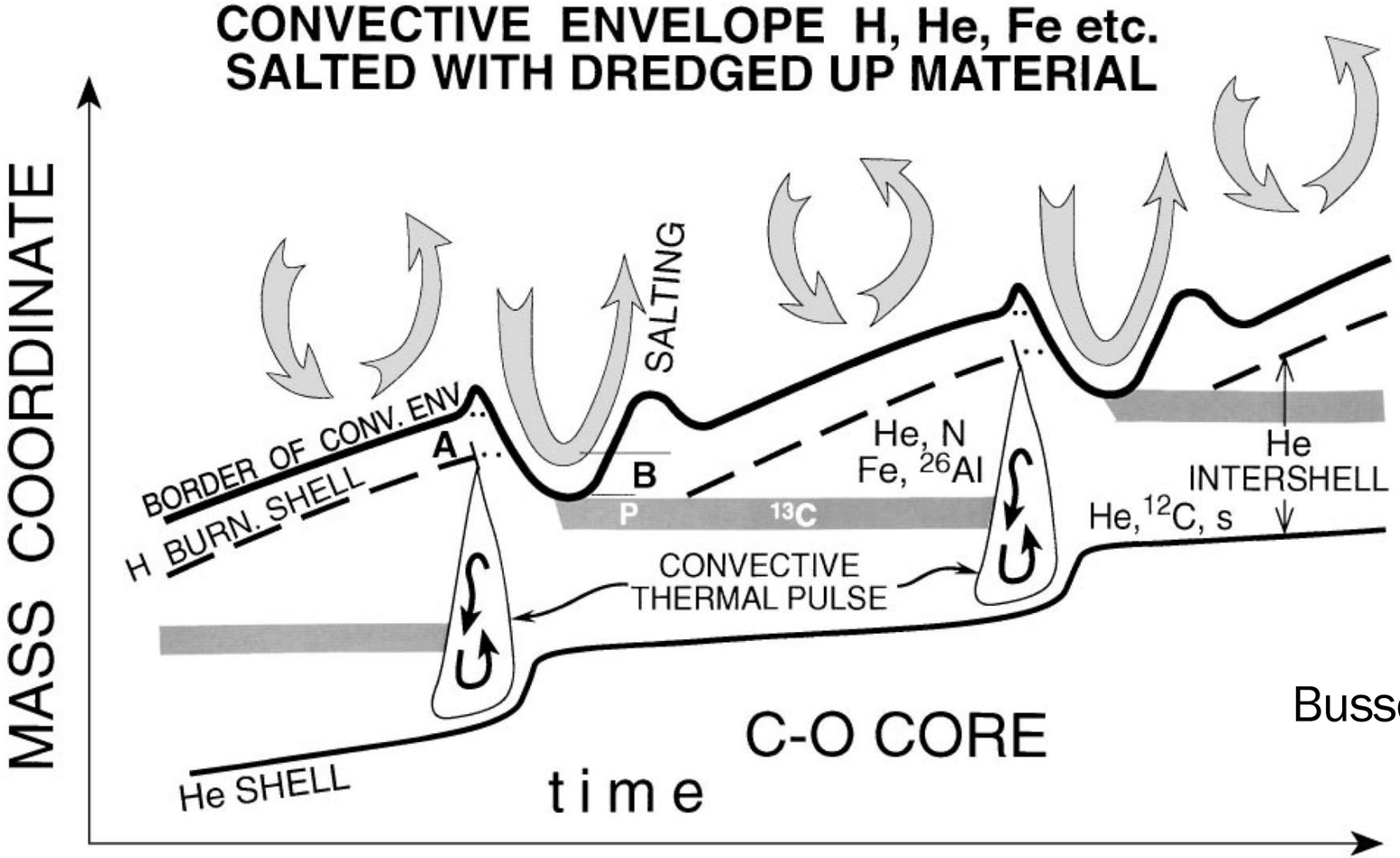


AGB phase

The sequence of thermal pulses has periods of roughly a few thousand years; this type of instability is associated with the variability observed in **Mira-type stars.**

During this phase, the **third dredge-up** occurs (with production of *s*-process elements).

AGB stars lose mass through intense stellar winds, reaching rates as high as 10^{-4} $M_\odot$ yr^{-1}.



Busso et al. 1999

Final phases for a star of $0.5<M<M_w$

M_w= **Limiting mass for the formation of a CO white dwarf**

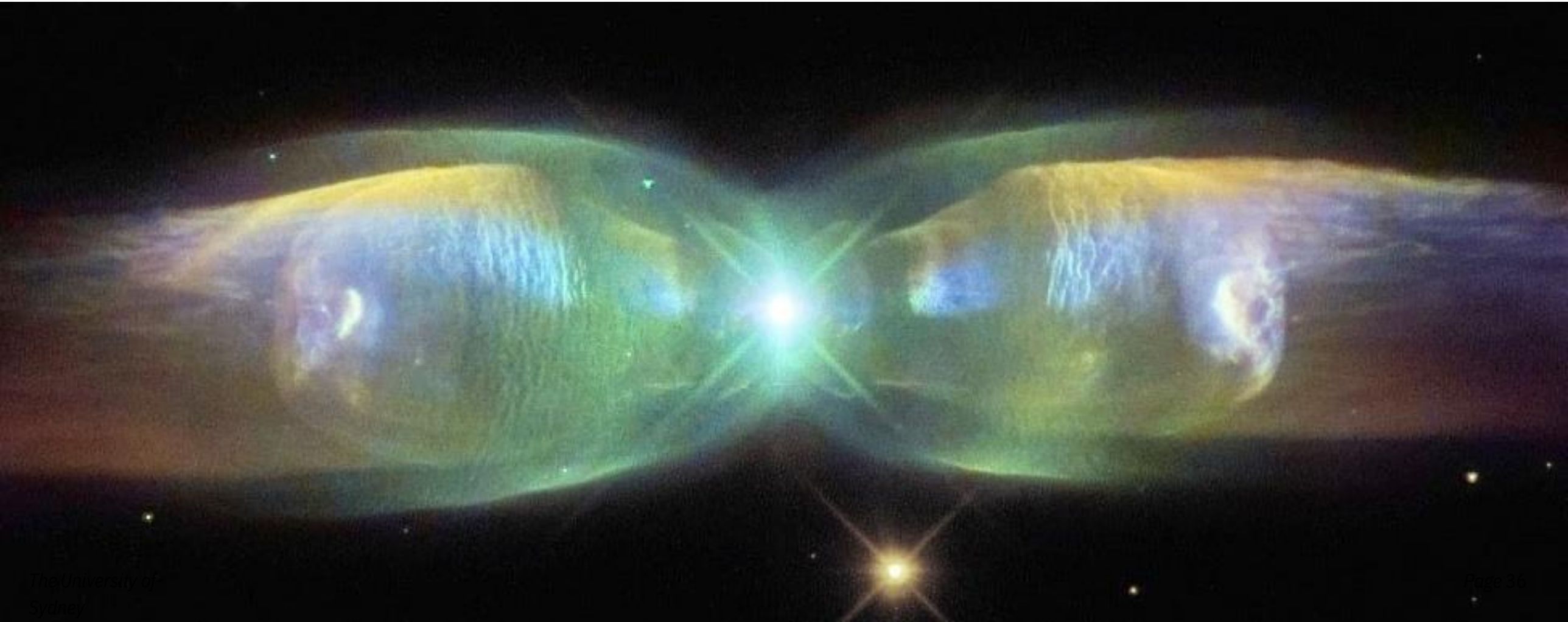
These stars contribute to the chemical enrichment of the Galaxy through stellar winds and planetary nebulae (releasing He, C^{12}, C^{13}, O^{17}, and s-process elements).
They end their lives as **carbon–oxygen white dwarfs**, with lifetimes ranging from well beyond the Hubble time to a few hundred million years.

Naively, we might expect the gas to be ejected spherically; the nebula, however, would appear as a ring because we see more gas along the sides of the bubble. This is the Helix nebula (NGC 7293), the nearest planetary nebula.

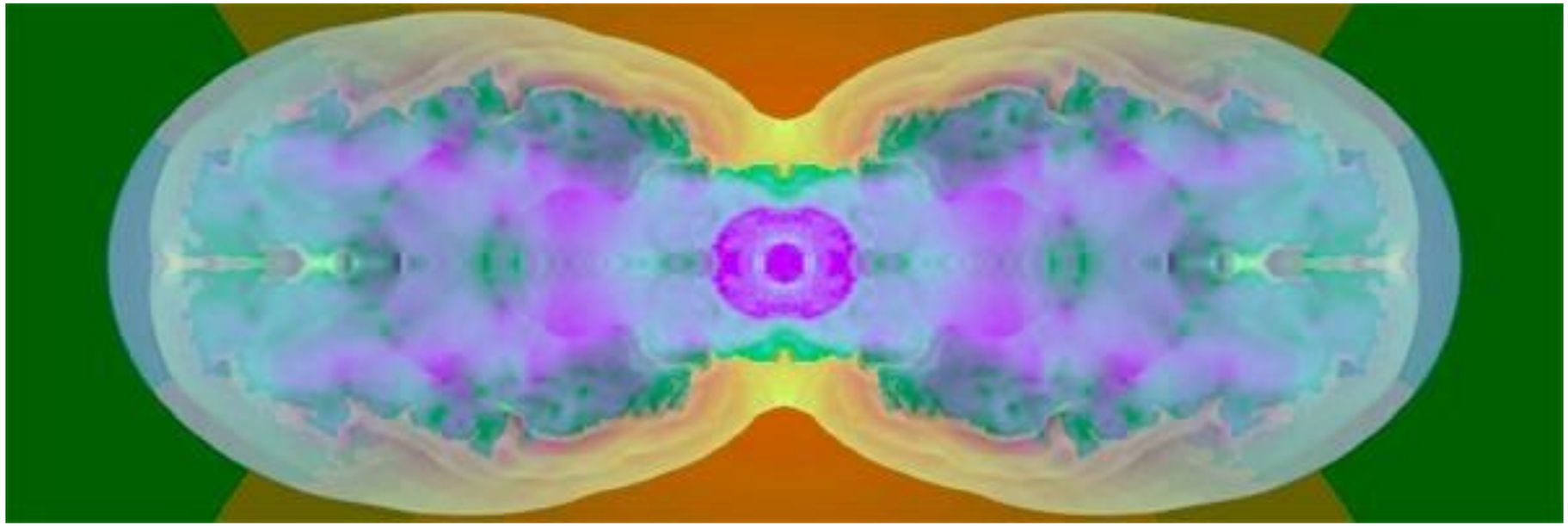
This animation shows the birth of the Helix Nebula. The star ejects its outer envelope at low velocity, exposing the hot core of the star. This star has a fast low-density wind that blows a big cavity in the dispersed envelope. UV radiation from the star heats the gas, causing it to glow.

However, many planetary nebulae have strange and wonderful shapes, indicating that the mass ejection process is very complicated. In binary systems, for instance, the gas outflow is influenced by the interactions between the stars. This is an HST/ X-ray picture of the Cat’s Eye nebula (NGC 6543).

The Butterfly nebula (M2-9) is very far from circular. The central object is a binary orbiting within a gaseous disk; the expelled envelope interacted with this disk to form the spectacular shape we see today. In fact, most planetary nebulae are bipolar.

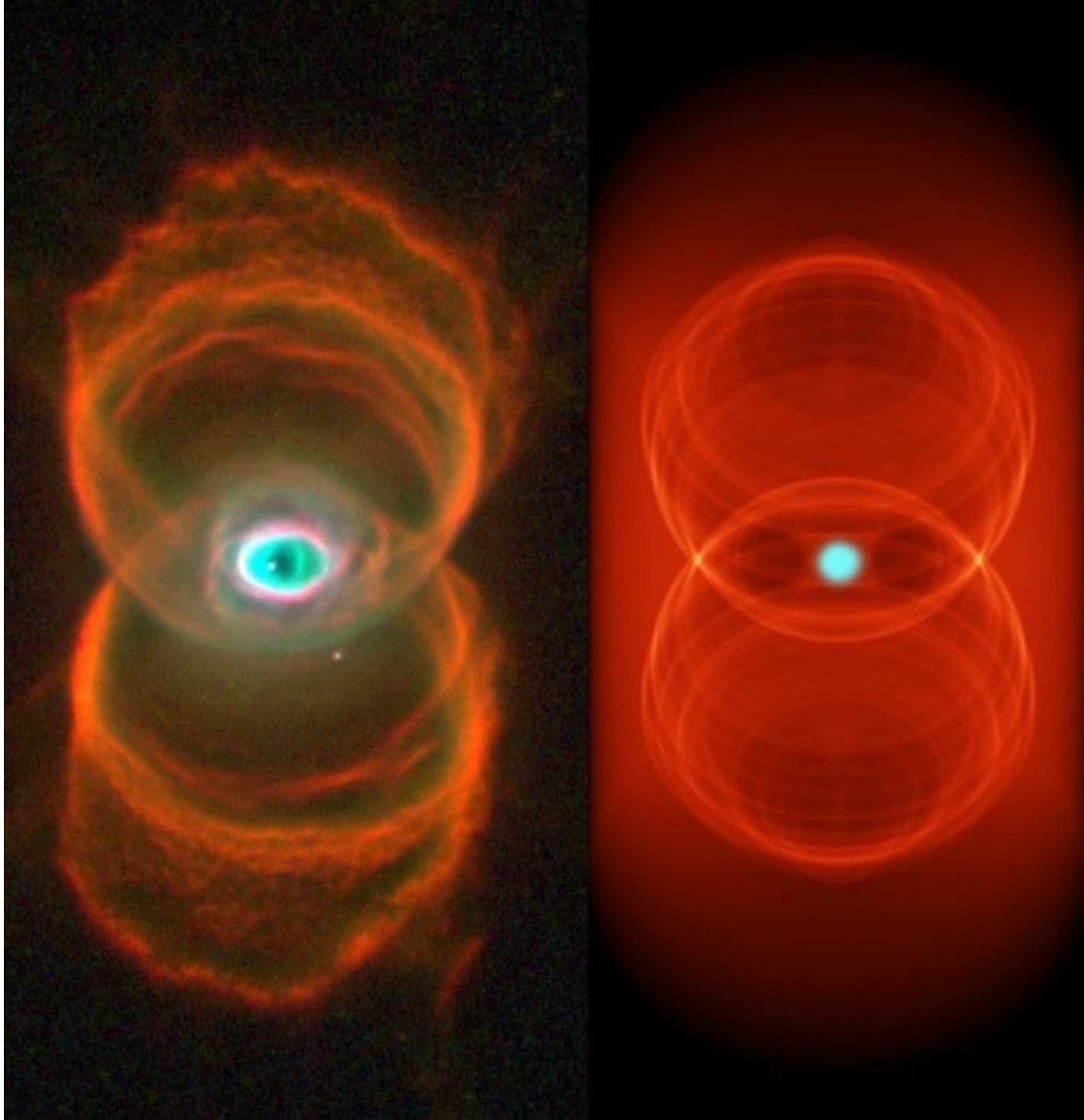


Simulations suggest that the strange shapes of planetary might be due to asymmetric winds from the progenitor confining the gas ejected during the planetary nebula phase. A disk-shaped cloud of gas around the star's equator forces the gas to escape towards the poles, forming a *bipolar nebula*.

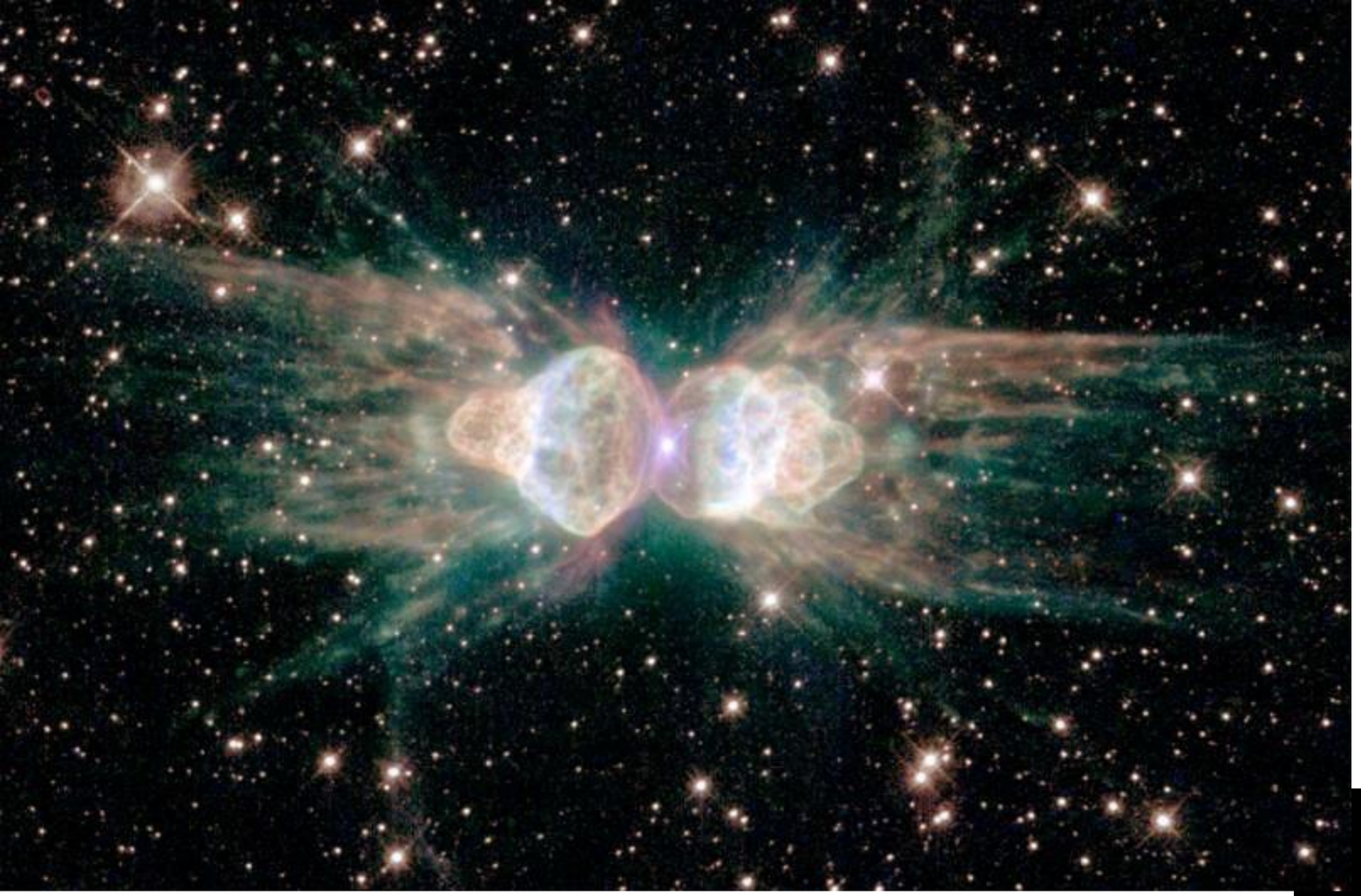


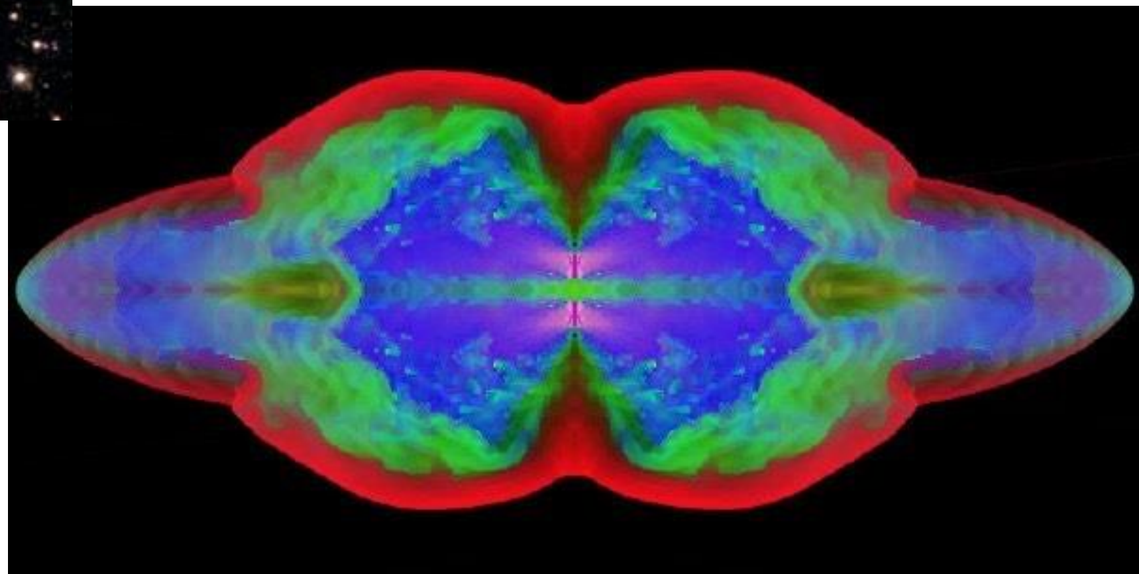
Simulation of gas escaping from a star, while being constrained by a torus of gas around the star.

These simulations reproduce some of the known planetary nebulae with amazing accuracy.



Hubble image of the “Hourglass nebula”, MyCn 18, compared to the interacting wind model viewed from an angle of 40o.

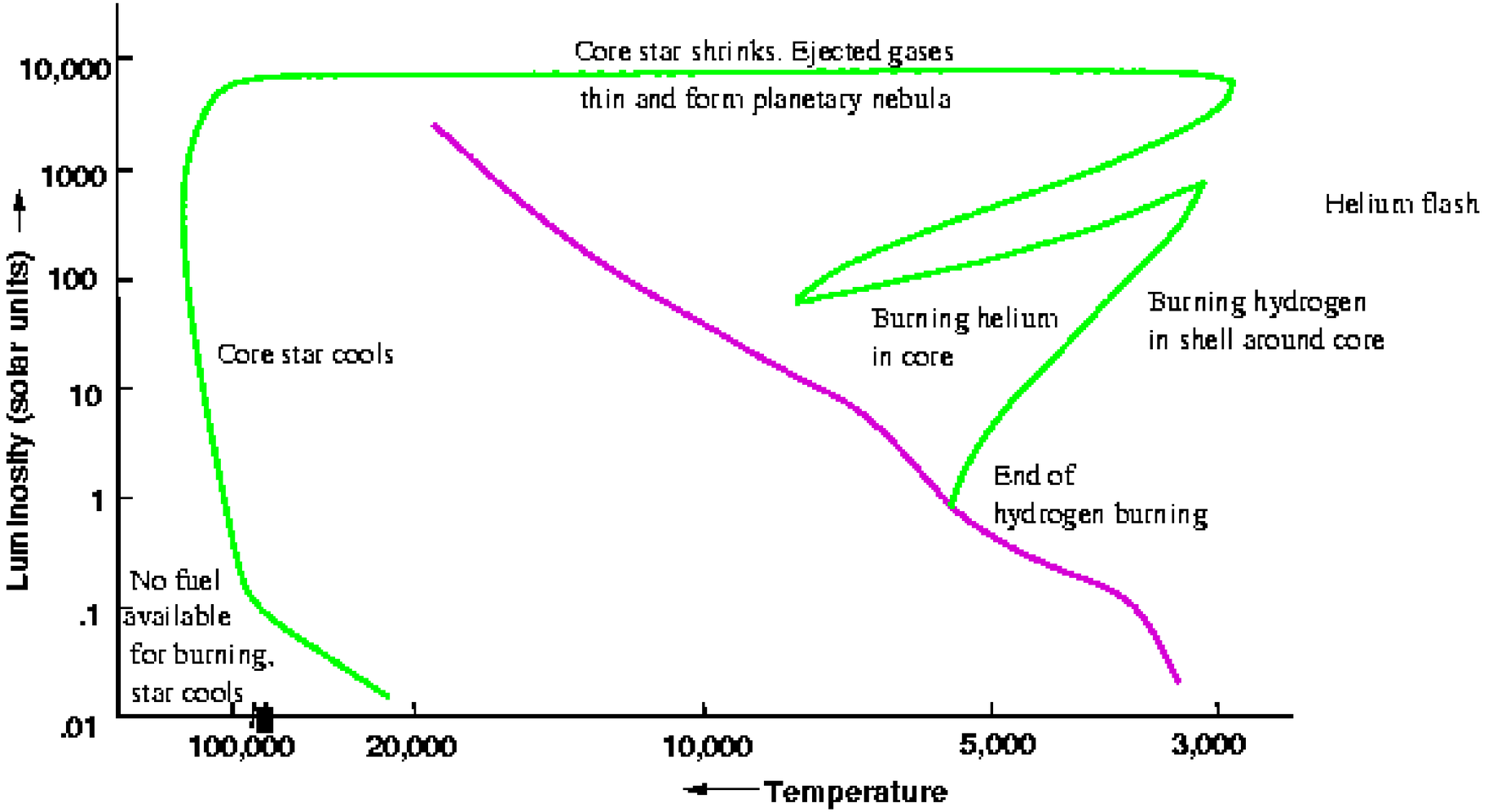




The “Ant Nebula”, Menzel 3.

M27: The Dumbbell Nebula

The remaining core of the star, consisting mostly of carbon and oxygen (ash from the helium burning), starts off very hot, but has no fuel source, i.e. no nuclear reactions are taking place to replace the energy lost by radiation. It will only cool and shrink: it has become a *white dwarf*.



The planetary nebula NGC 2440, with its central white dwarf, one of the hottest white dwarfs known.

The planetary nebula known as the “Skull Nebula”, NGC 246, with its white dwarf, the fainter member of the binary star system seen at the nebula's centre.

The evidence for stellar evolution

We’ve built up a wonderful picture for stellar evolution: but the timescales are so long that we can never observe a star evolving. How certain are we about all this?

The best evidence comes from studies of *star clusters*. Clusters of stars are ubiquitous in the Galaxy.

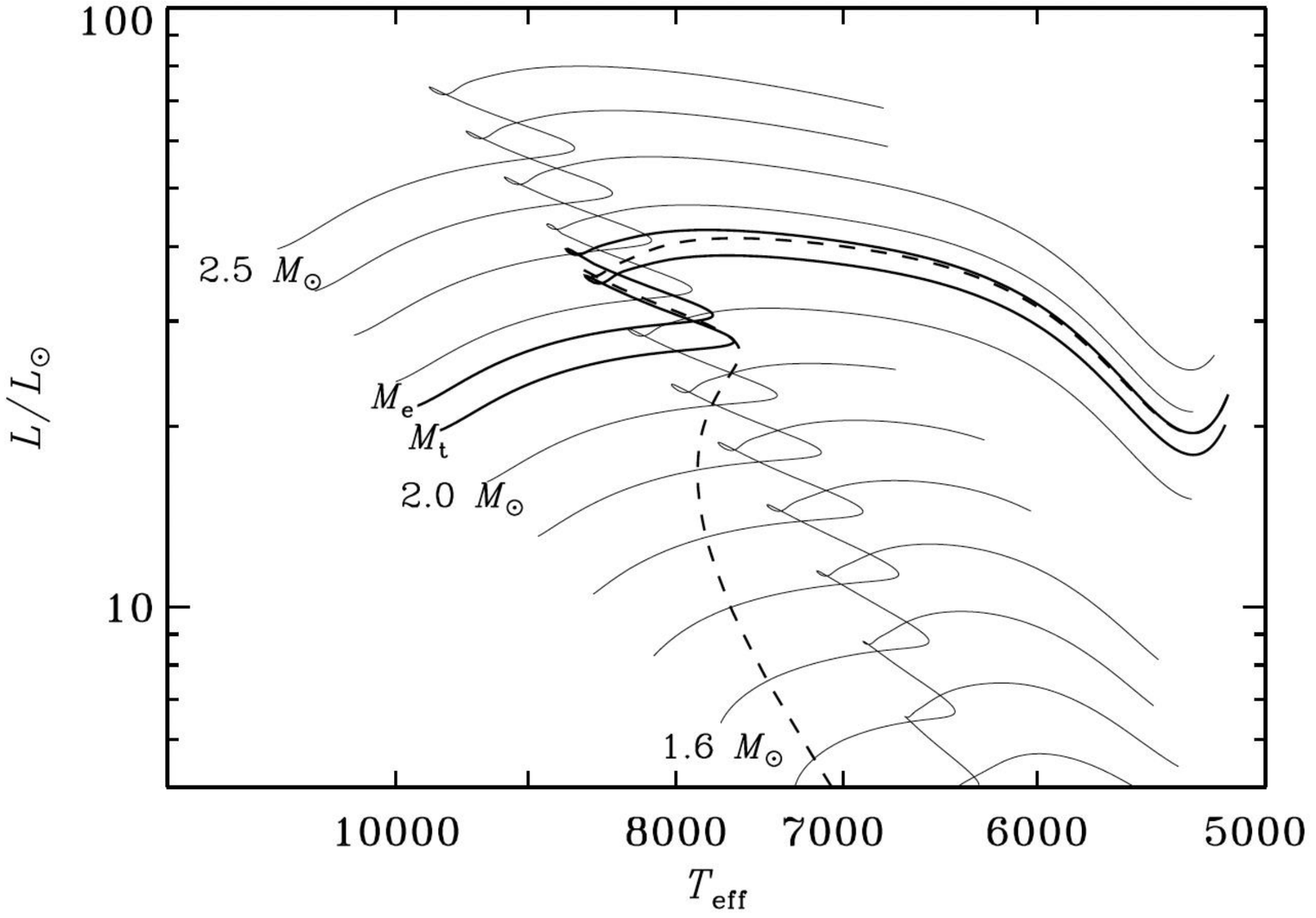
Generally, we divide clusters into two classes: *open clusters*, consisting of 100–1000 young stars, all born from the same molecular cloud;

The Jewel Box Cluster (NGC 4755, or Kappa Crucis)

...and *globular clusters*, which are extremely old groups of 10,000–1,000,000 stars, born in the very earliest stages of the Galaxy's evolution.

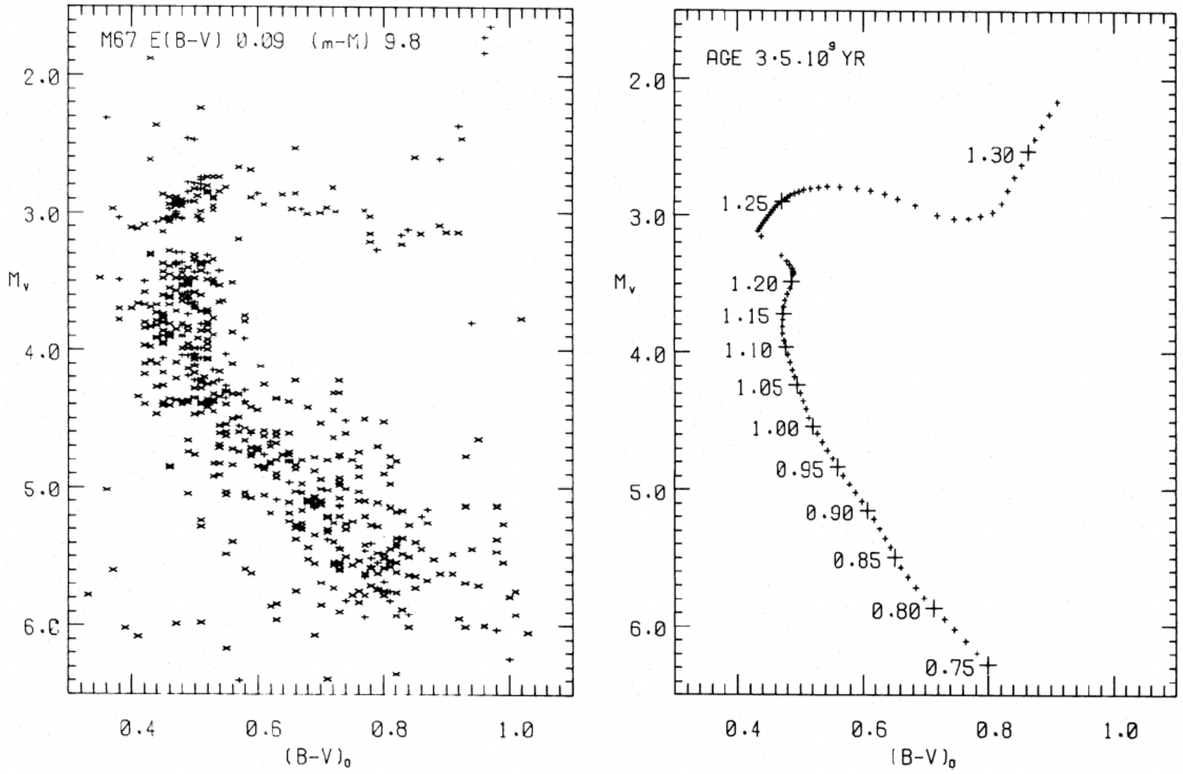
The dense globular cluster M80(or NGC 6093)

Isochrones and evolutionary tracks

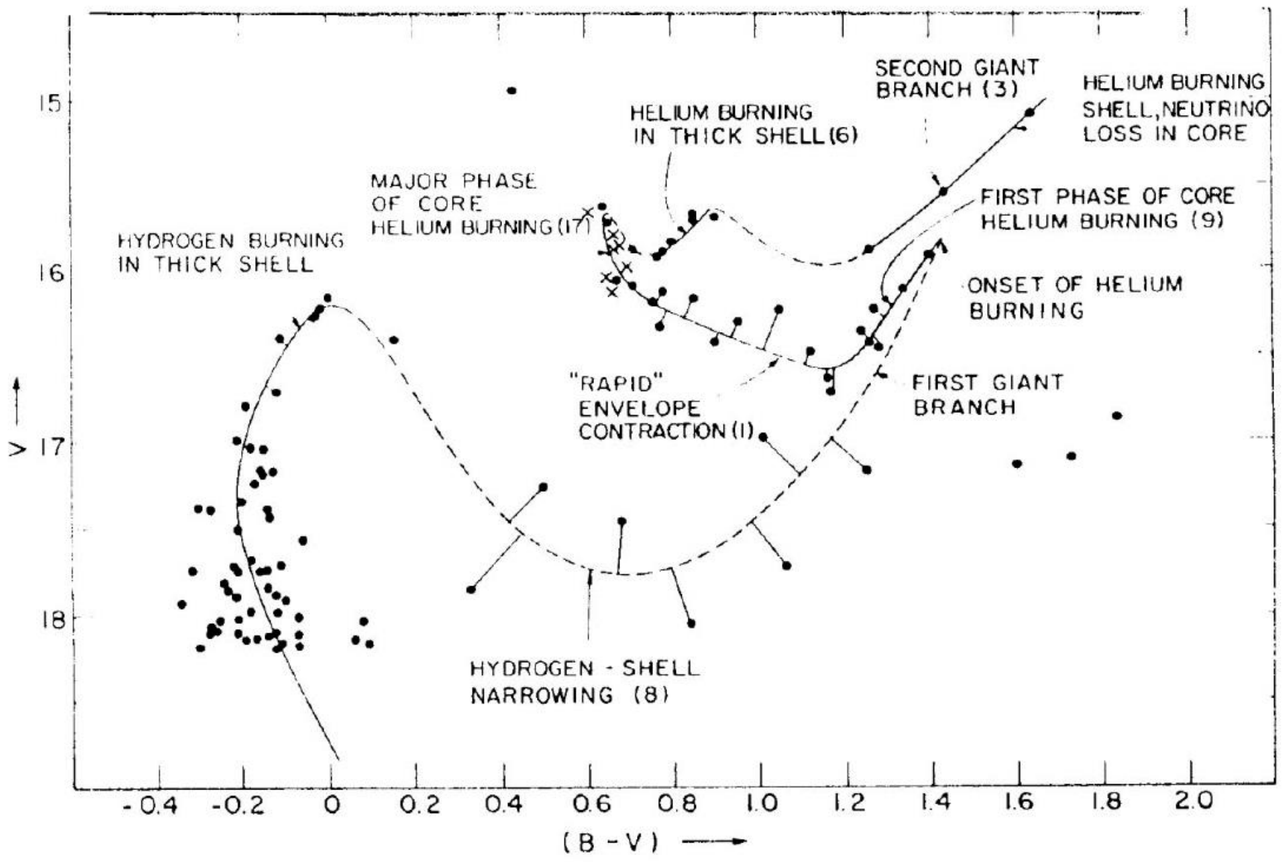


An isochrone (dashed line) lies between the evolutionary tracks for the turn-off mass Mt and the end-point mass Me, indicated by heavy continuous lines;
here we only consider evolution to the base of the red giant branch. The age of the isochrone is 8 10^8 years, corresponding to Mt = 2.1Msun and Me = 2.15Msun .
Additional evolutionary tracks are shown with thin continuous lines, for models with
a step in mass of 0.1Msun

Results for an open cluster

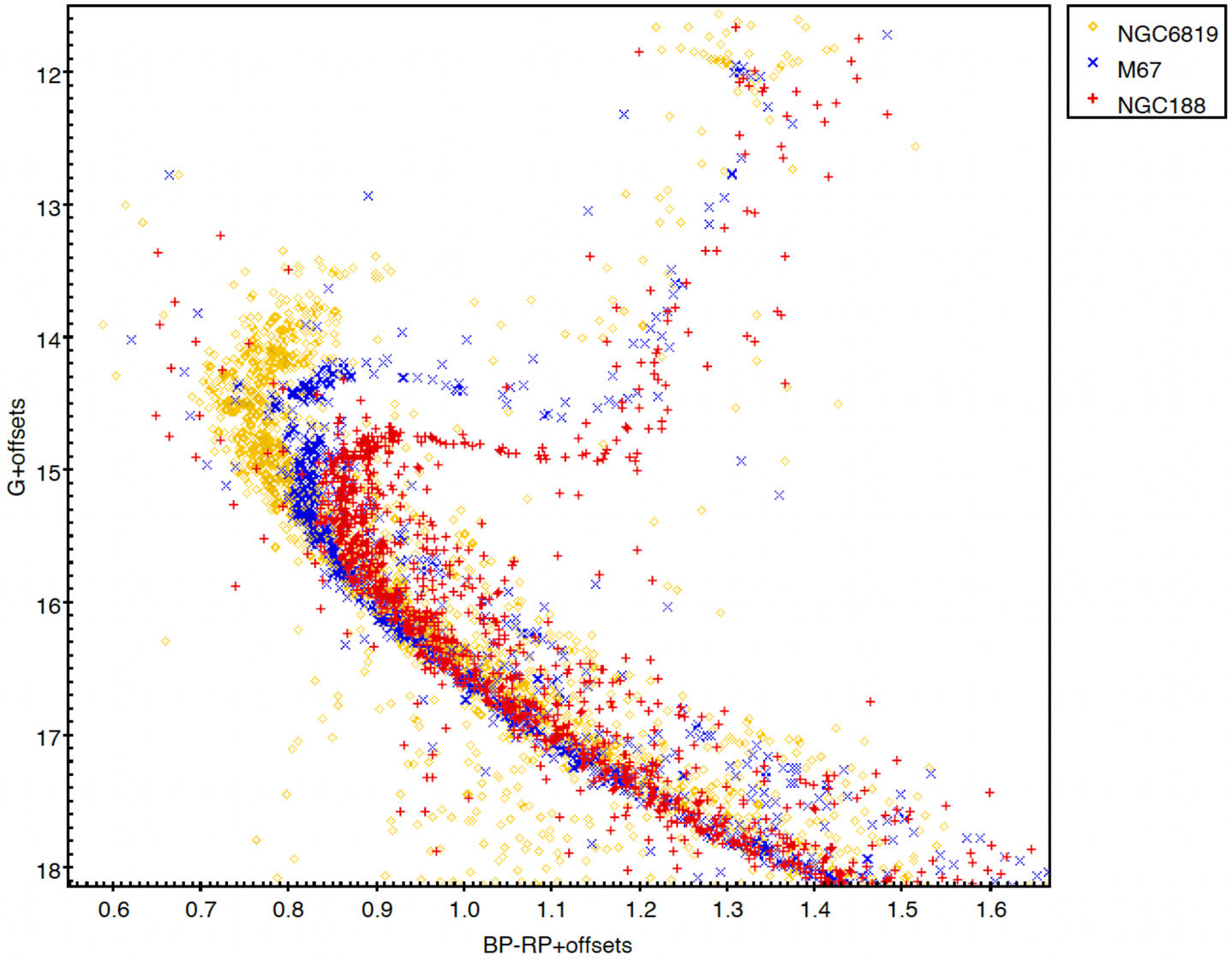


Observed HR diagram for the open cluster Messier 67 (left) and theoretical isochrone for an age of 3.5×10^9 years (right); the theoretical values for $T_{\rm eff}$ and $L_{\rm s}$ have been translated into B-V and absolute visual magnitude M_v to allow a direct comparison with the observations. On the isochrone the masses have been indicated in units of the solar mass. In general the agreement between theory and observation is quite good. (From Morgan & Eggleton 1978).

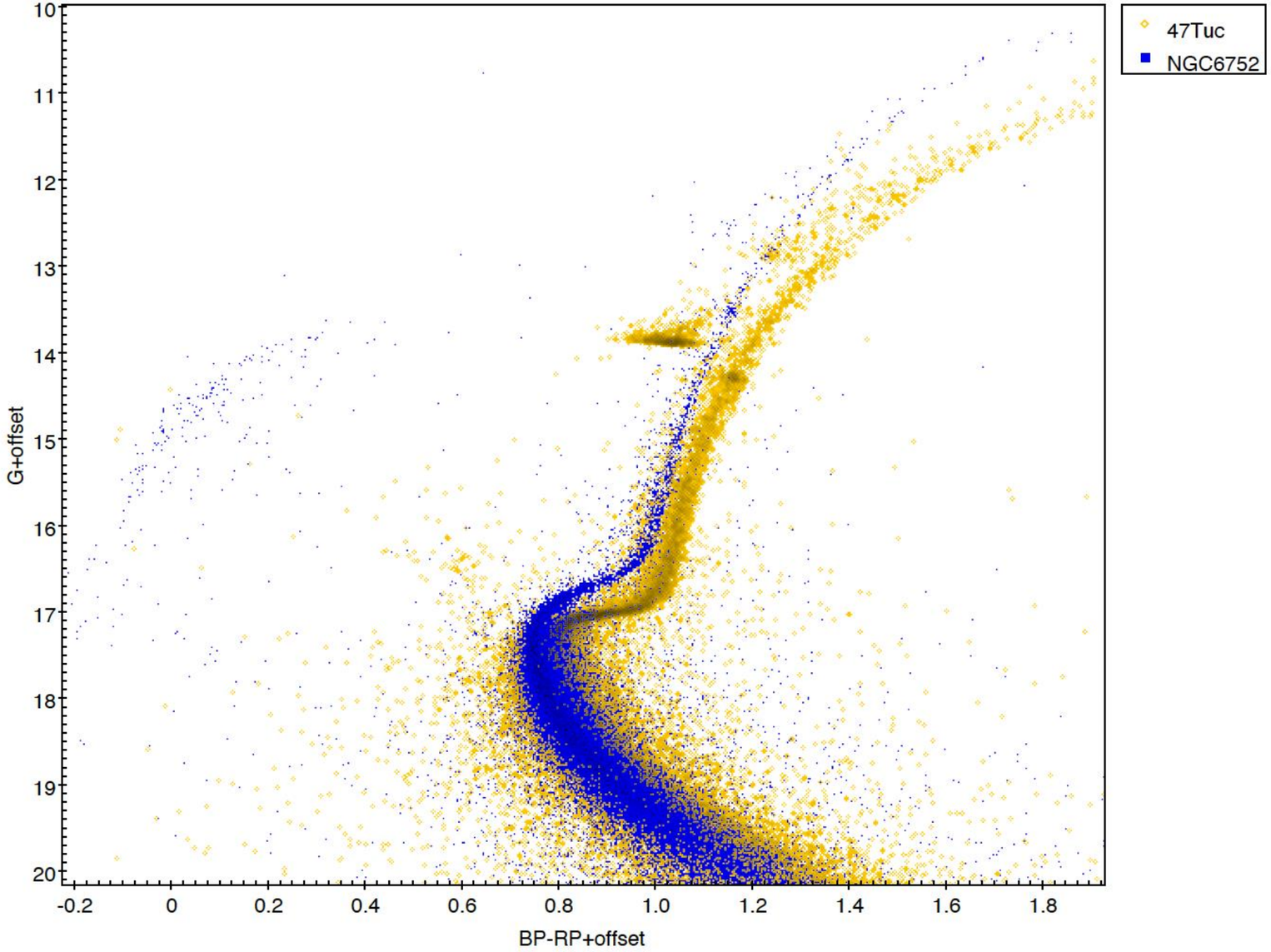


Observed HR diagram for a stellar cluster (NGC 1866 in the Large Magellanic Cloud). In the figure are indicated evolutionary phases found by comparing with the theoretical evolution track shown in Figure 12.2. The numbers in brackets indicate the number of stars in each phase of evolution. (From Iben 1985).

Newer results with GAIA



What is the difference?



The mass range of 8 -12 Msun

Stars with $M_w < M < M_{\mathrm{up}}$,where $M_{\mathrm{up}}$is defined as the upper mass limit for forming a degenerate C–O core, and its value depends on the treatment of convection and mass loss.

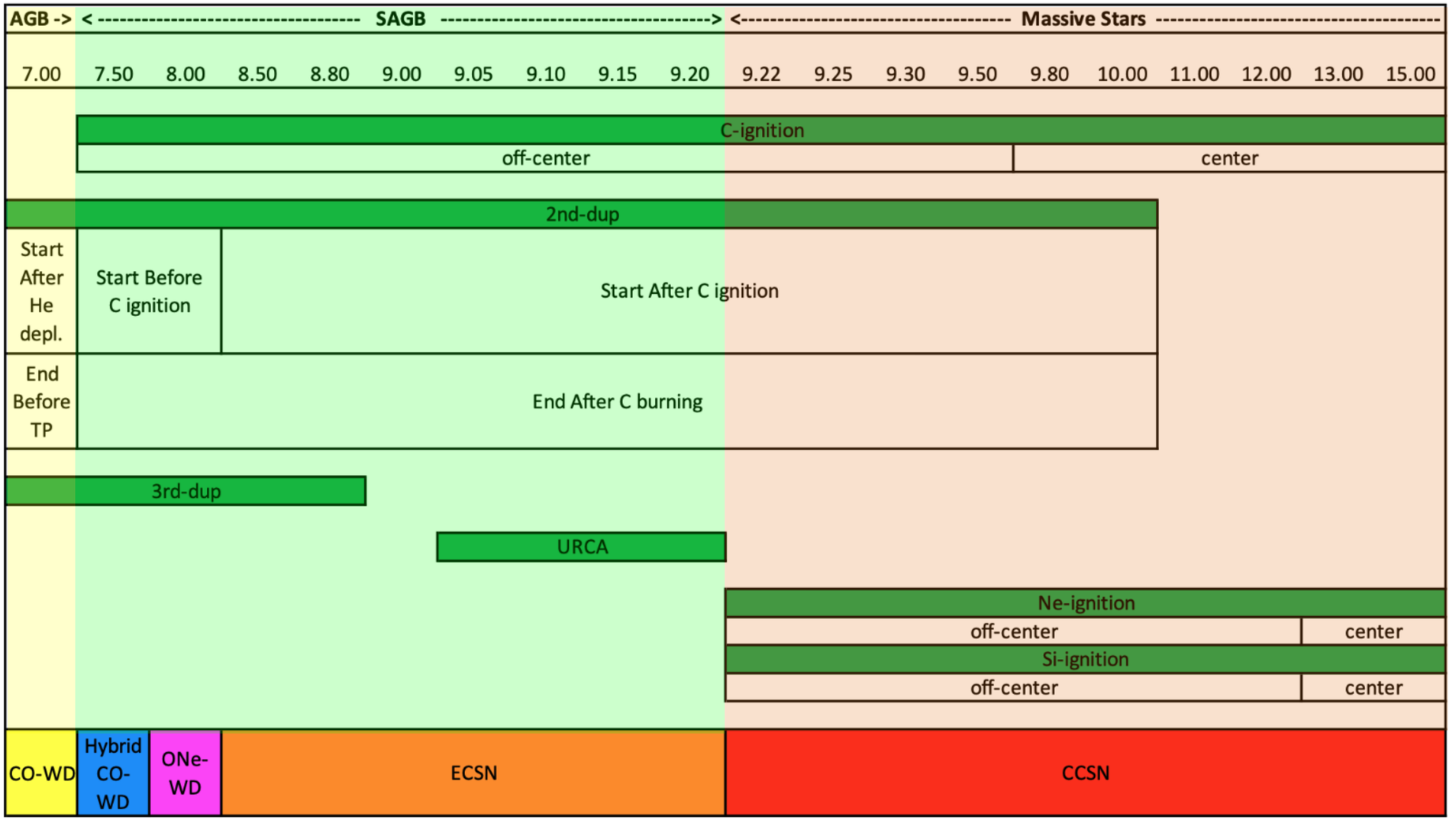
Unless $M_w \approx M_{\mathrm{up}}$,these stars should ignite carbon in a highly degenerate core when the core mass reaches the **Chandrasekhar mass** M_{Ch} ,resulting in a catastrophic explosion that would produce a supernova.
Such supernovae, although not yet observed, would be called **Type I 1/2 supernovae**, because their characteristics would be intermediate between Type I and Type II supernovae (see the chapter on supernovae).

Observations, however, suggest $M_w \approx 8M_\odot$

In this case, such supernovae would not exist because stellar models indicate that M_{up}should be at most $8M_\odot$ .The precise value of this mass again depends on the stellar models. Models including **overshooting** suggest $M_{\mathrm{up}} = 5\text{–}6M_\odot$.

If this is indeed the case, we must conclude that all stars up to M_{up}or $M_w$die as **C–O white dwarfs**, meaning that none of them reach a core mass sufficient to ignite carbon.

The lifetimes of these stars range from billions to tens of millions of years, and they contribute to **galactic enrichment** through **stellar winds** and **planetary nebulae**.



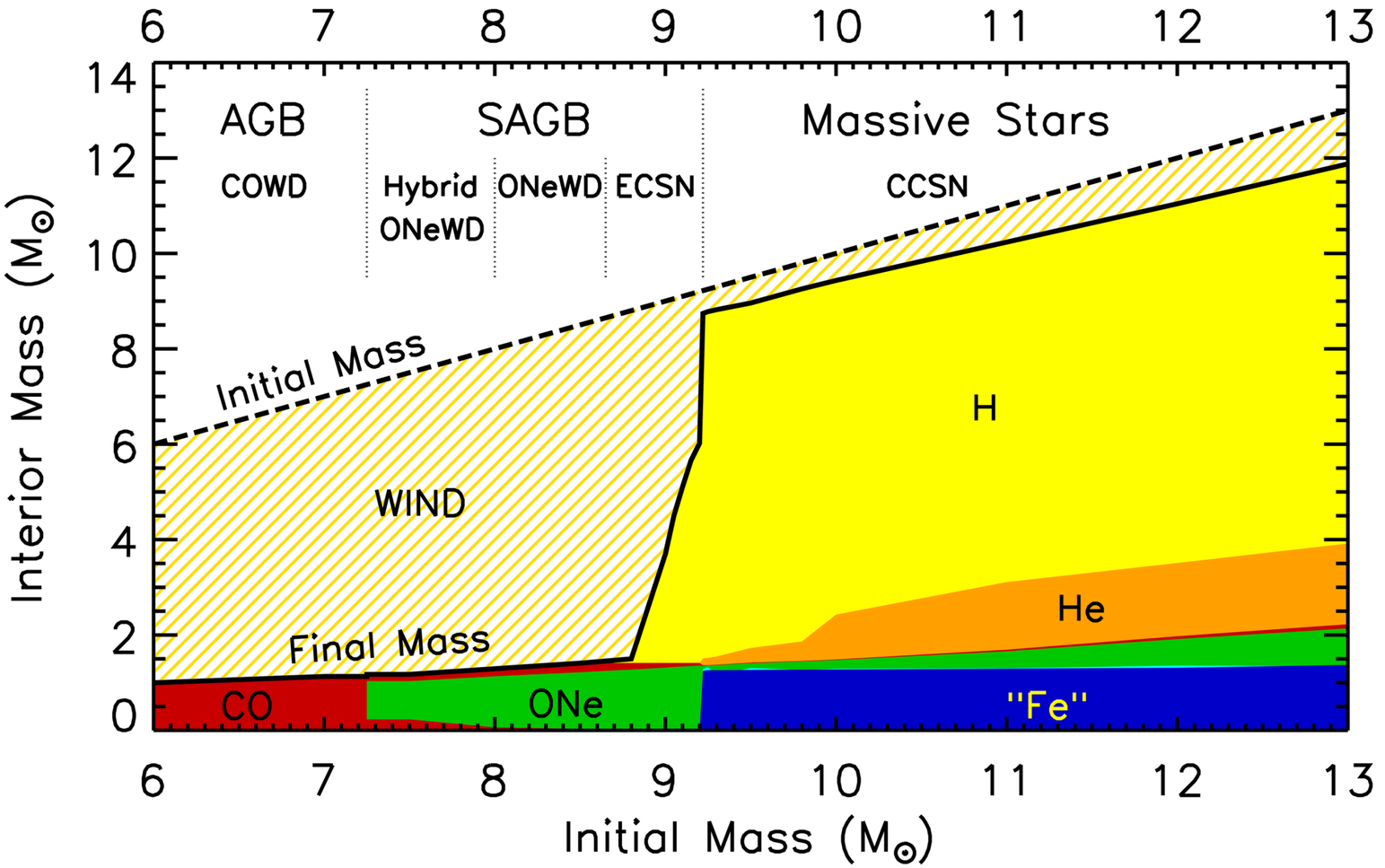
Limongi, Roberti & Chieffi (2024)

The ejected masses decrease substantially for stars with M<13 $M_\odot$ – Extrapolation dangerous

The high number of stars in the mass interval 9.22-13 $M_\odot$ is compensated by the small amount of ejected masses

Diversity of the optical display both luminosity and duration of the plateau

Limongi, Roberti & Chieffi (2024)



The ejected masses decrease substantially for stars with M<13 $M_\odot$ – Extrapolation dangerous

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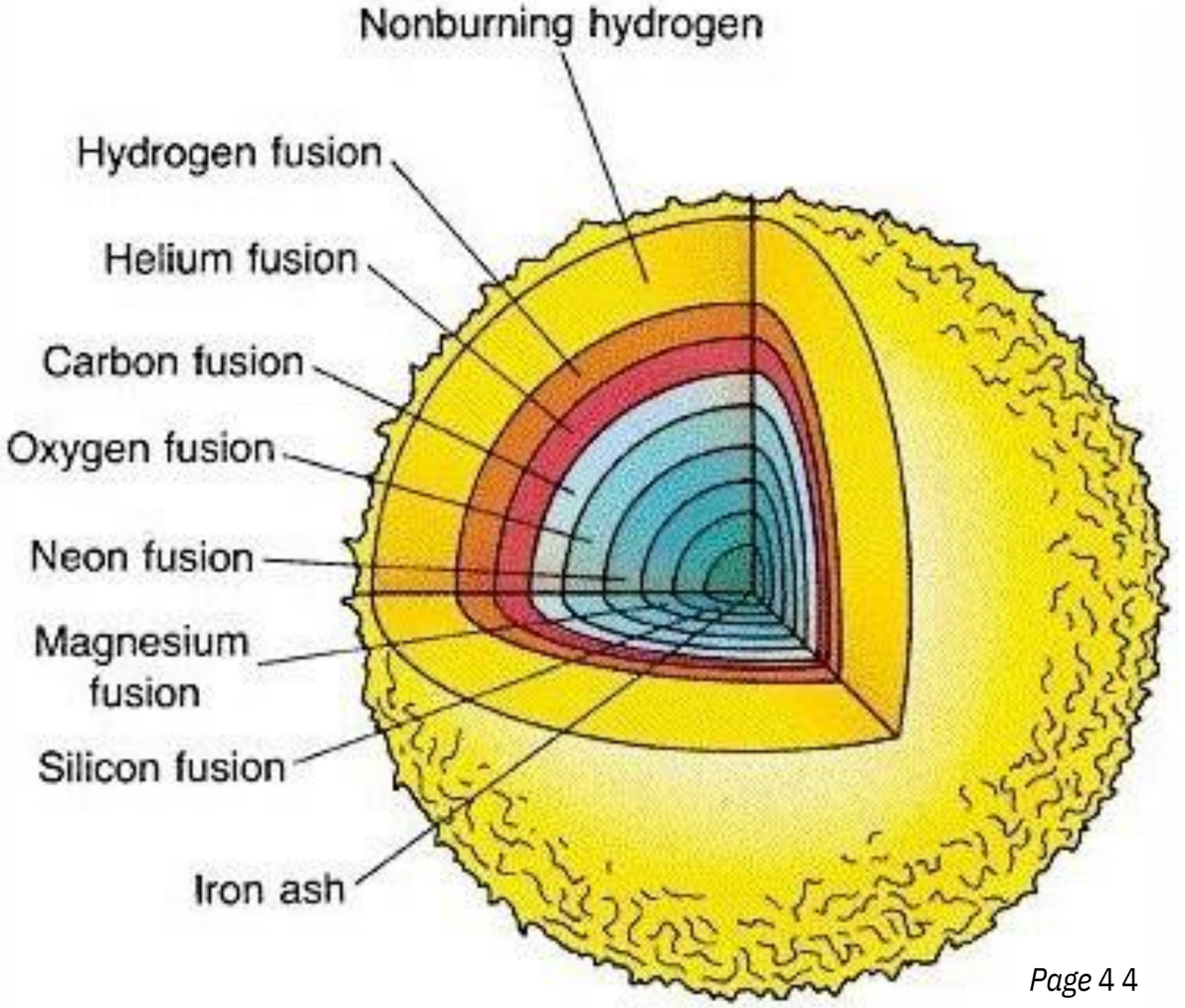
Diversity of the optical display both luminosity and duration of the plateau

Massive stars > 12Msun

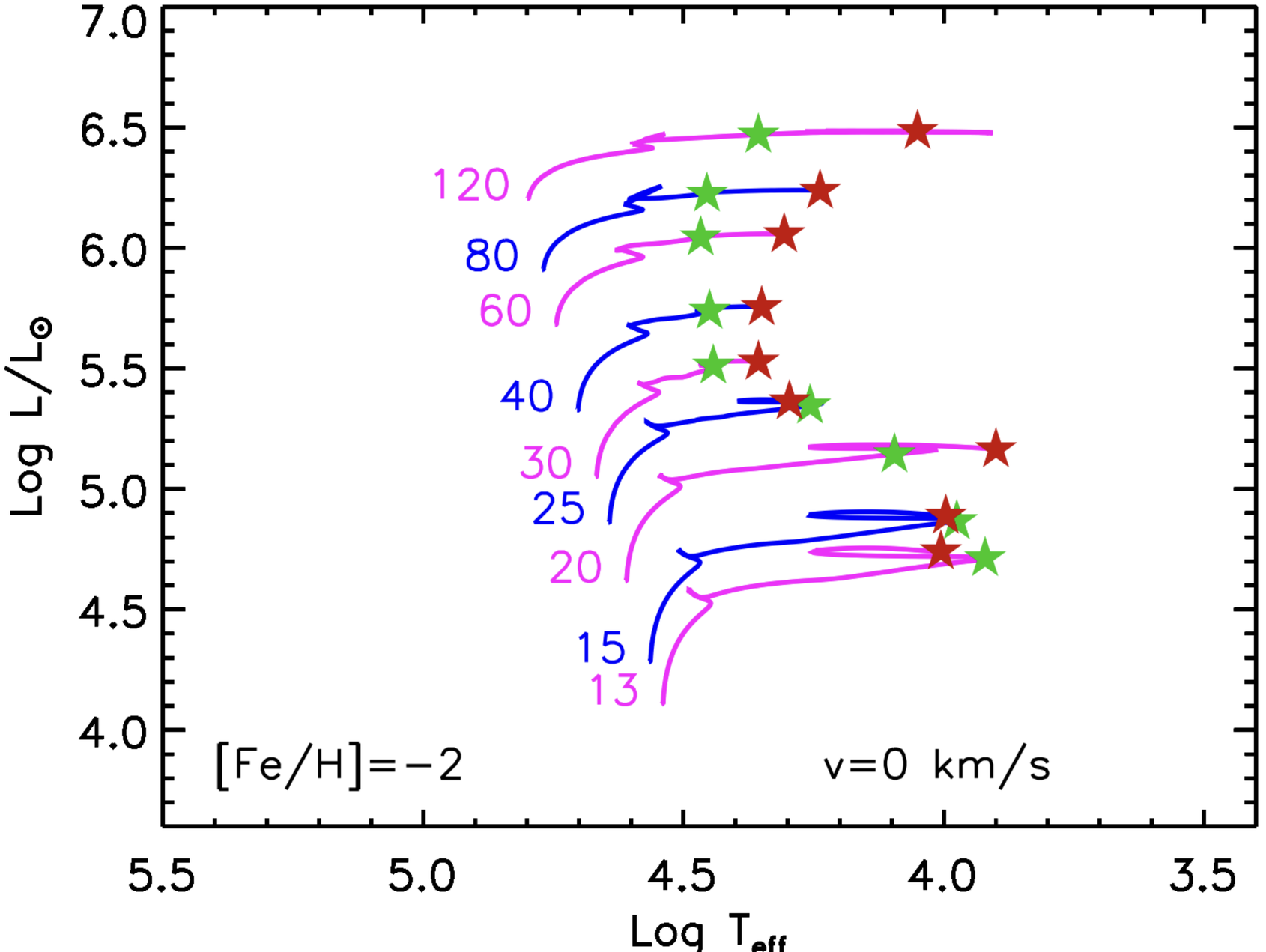
High mass stars have much more violent endings. Massive stars can reach temperatures in their core high enough to begin fusing carbon, producing neon and oxygen. When the carbon is exhausted in the core, it contracts and carbon ignites in a shell. This pattern of core ignition and shell ignition continues with fuel after fuel, until the star develops a layered structure.



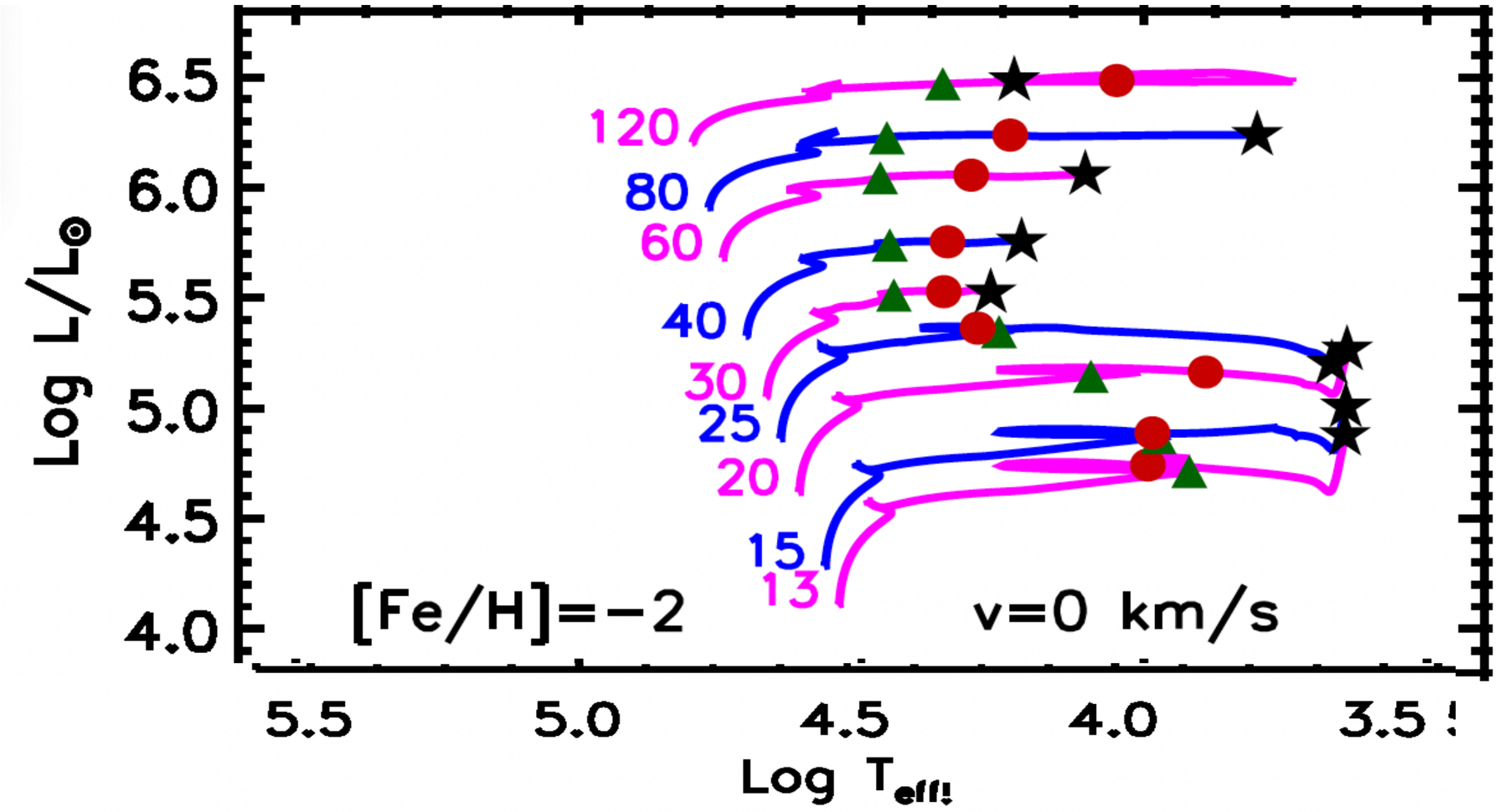
Final stages of massive supergiant stars (M > 10 M⊙)

- Very high mass and luminosity.
- Rapid evolution, with no “loops” in the H–R diagram.
- After the main sequence, the star expands and moves once to the right, becoming a red supergiant.
- Successive core fusion stages produce heavier elements up to iron.
- The iron core collapses, triggering a core-collapse supernova (Type II), leaving a neutron star or a black hole.

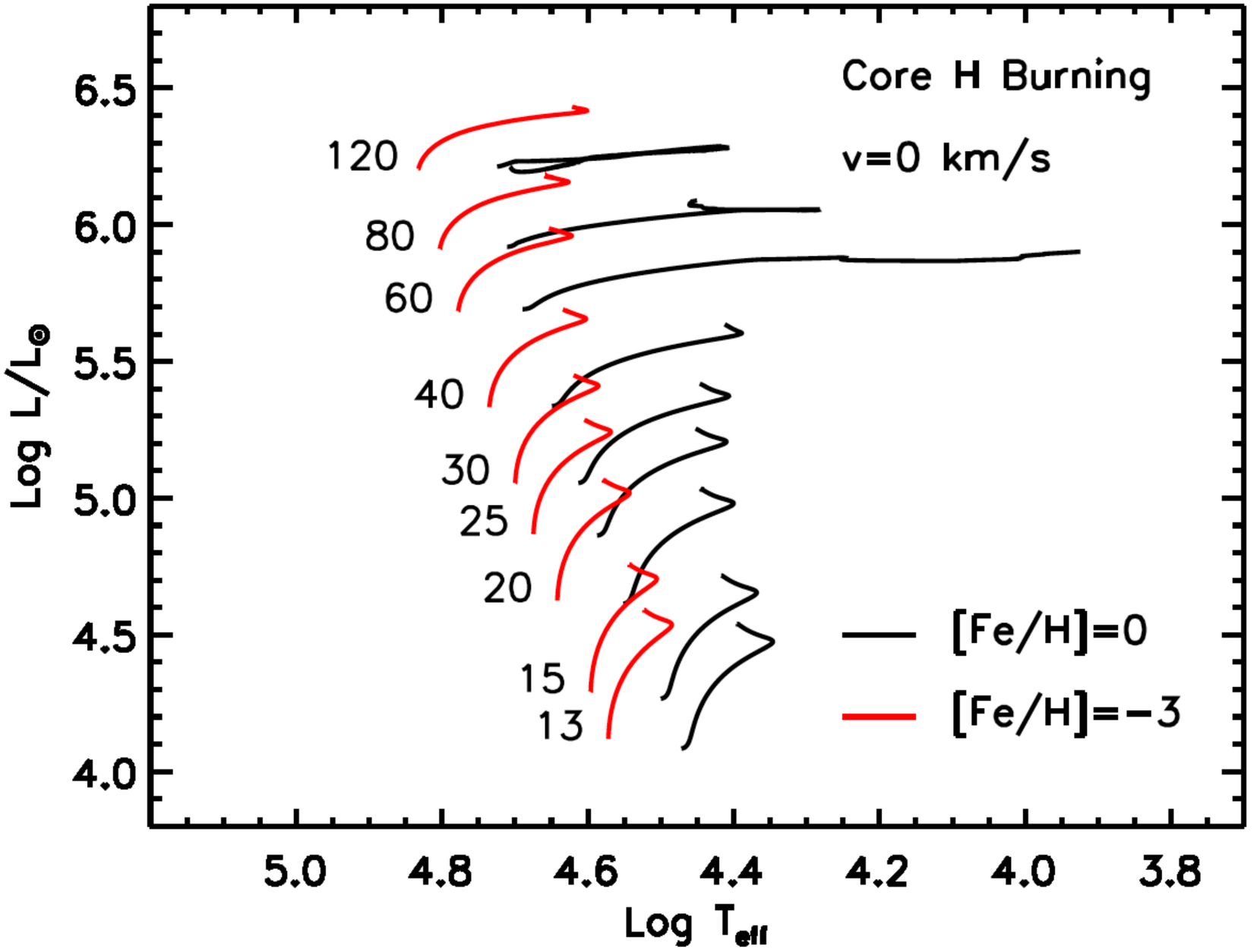
Limongi and Chieffi 2018



The various symbols mark the central He-ignition (green triangles), the central He-exhaustion (red dots), and the final position at the presupernova stage (black star)



The various symbols mark the central He-ignition (green triangles), the central He-exhaustion (red dots), and the final position at the presupernova stage (black star)


Core H Burning
v=0 km/s
120
80
60
40
30
25
20
15
13
[Fe/H]=0
[Fe/H]=−3
Log L/L⊙
Log T_eff
6.5
6.0
5.5
5.0
4.5
4.0
5.0
4.8
4.6
4.4
4.2
4.0
3.8

Final stages of massive supergiant stars (M > 10 M⊙)

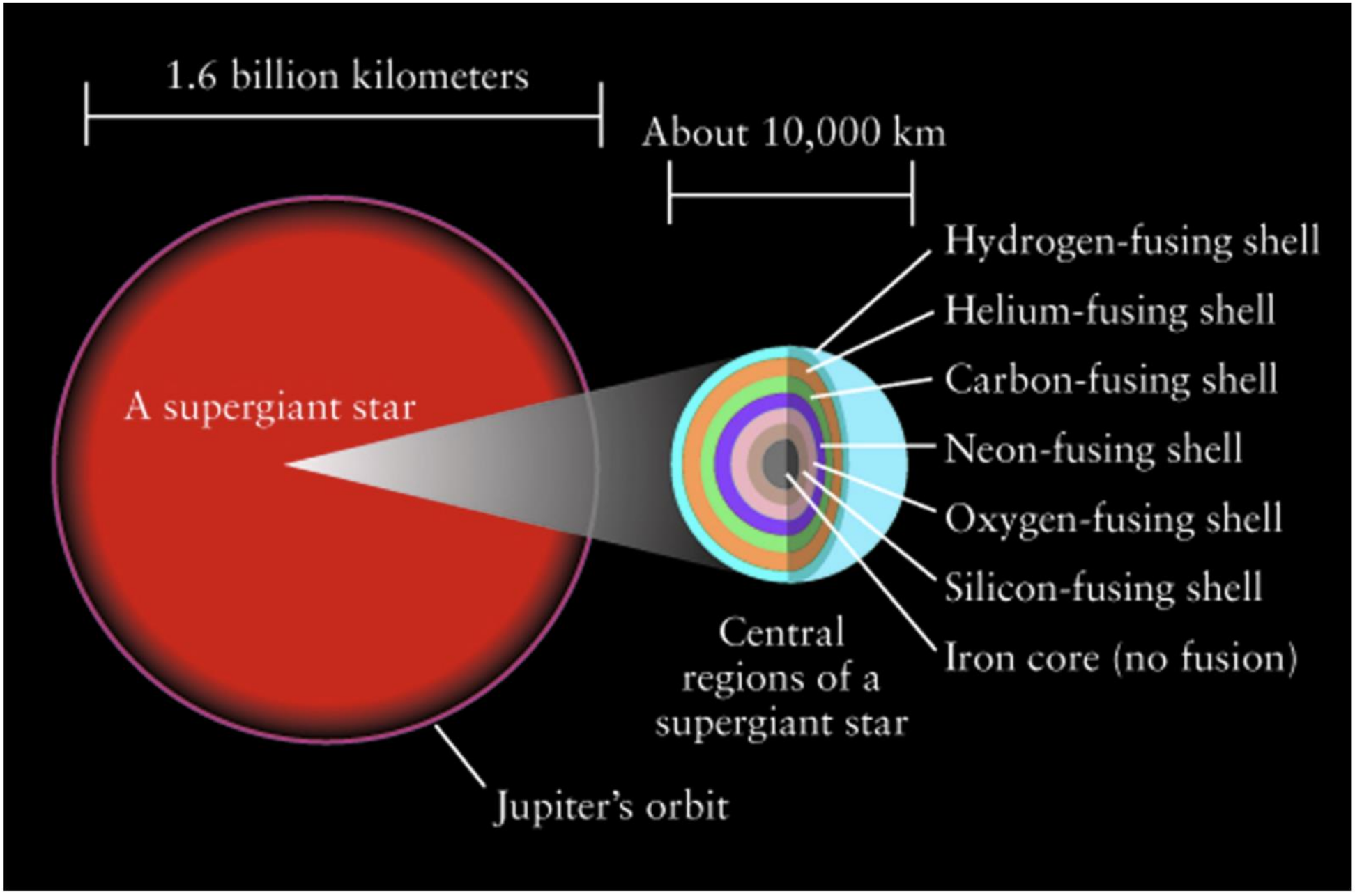
Timescales for the subsequent burining phases for a 25Msun

Fuel	Temperature	Time (y)	Sample reaction
Hydrogen	(K) 40 million	7,000,000	$4\,^{1}H \rightarrow {}^{4}He$
Helium	200 million	500,000	$3\,^{4}He \rightarrow {}^{12}C$
Carbon	600 million	600	$2\,^{12}C \rightarrow {}^{20}Ne + {}^{4}He$
Neon	1.2 billion	1	$^{20}Ne + {}^{4}He \rightarrow {}^{16}O + 2\,^{4}He$
Oxygen	1.5 billion	0.5	$2\,^{16}O \rightarrow {}^{28}Si + {}^{4}He$
Silicon	2.7 billion	1 day	$2\,^{28}Si \rightarrow {}^{56}Fe$

Advanced Nuclear Burning Stages
(e.g., 20 solar masses)

Fuel	Main Product	Secondary Products	Temp (10^9 K)	Time (yr)
H	He	^{14}N	0.02	10^7
He	C,O	^{18}O,^{22}Ne s- process	0.2	10^6
C	Ne, Mg	Na	0.8	10^3
Ne	O, Mg	Al, P	1.5	3
O	Si, S	Cl, Ar K, Ca	2.0	0.8
Si	Fe	Ti, V, Cr Mn, Co, Ni	3.5	1 week

(Credit: S. Woosley, https://www.ucolick.org/ woosley/ay112-14/lectures/


1.6 billion kilometers
About 10,000 km
A supergiant star
Hydrogen-fusing shell
Helium-fusing shell
Carbon-fusing shell
Neon-fusing shell
Oxygen-fusing shell
Silicon-fusing shell
Iron core (no fusion)
Central regions of a supergiant star
Jupiter's orbit

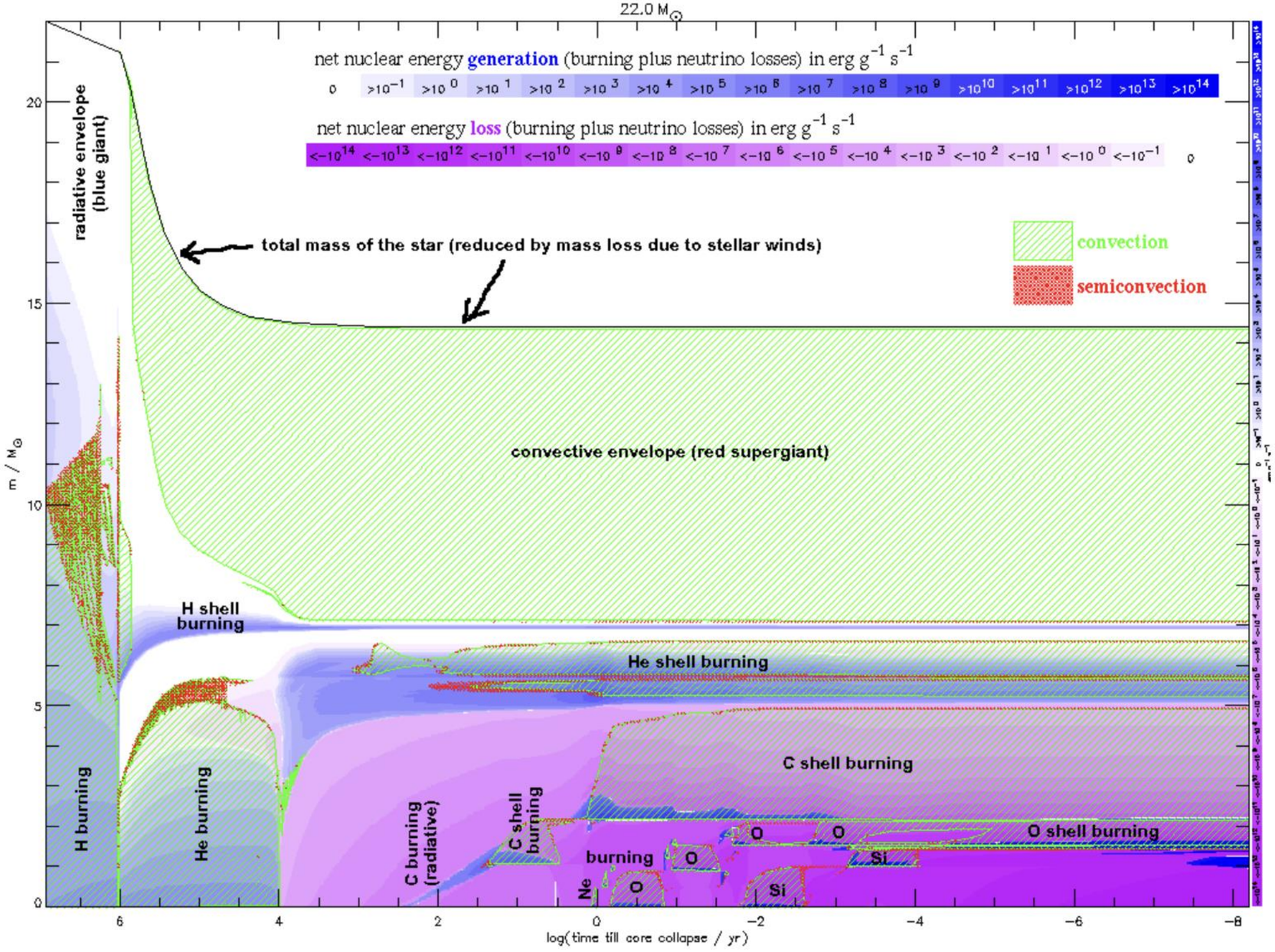
The evolution and explosion of massive stars

S. E. Woosley, A. Heger, and T. A. Weaver
Rev. Mod. Phys. **74**, 1015 – Published 7 November 2002

«Like all true stars, massive stars are gravitationally confined thermonuclear reactors whose composition evolves as energy is lost to radiation and neutrinos. Unlike lower-mass stars (M<~8Msolar), however, no point is ever reached at which a massive star can be fully supported by electron degeneracy. Instead, the center evolves to ever higher temperatures, fusing ever heavier elements until a core of iron is produced. The collapse of this iron core to a neutron star releases an enormous amount of energy, a tiny fraction of which is sufficient to explode the star as a supernova. ...»

During the later stages of burning, energy is lost to neutrinos: most of the star's energy, by the time carbon burning starts.

Fuel	Temperature(K)	Time (y)	photon luminosity (J/s)	neutrino luminosity (J/s)
Hydrogen	40 million	7,000,000	2.7×10^{31}	–
Helium	200 million	500,000	5.3×10^{31}	$<1.0 \times 10^{29}$
Carbon	600 million	600	4.3×10^{31}	7.4×10^{32}
Neon	1.2 billion	1	4.4×10^{31}	1.2×10^{36}
Oxygen	1.5 billion	0.5	4.4×10^{31}	7.4×10^{36}
Silicon	2.7 billion	1 day	4.4×10^{31}	3.1×10^{38}



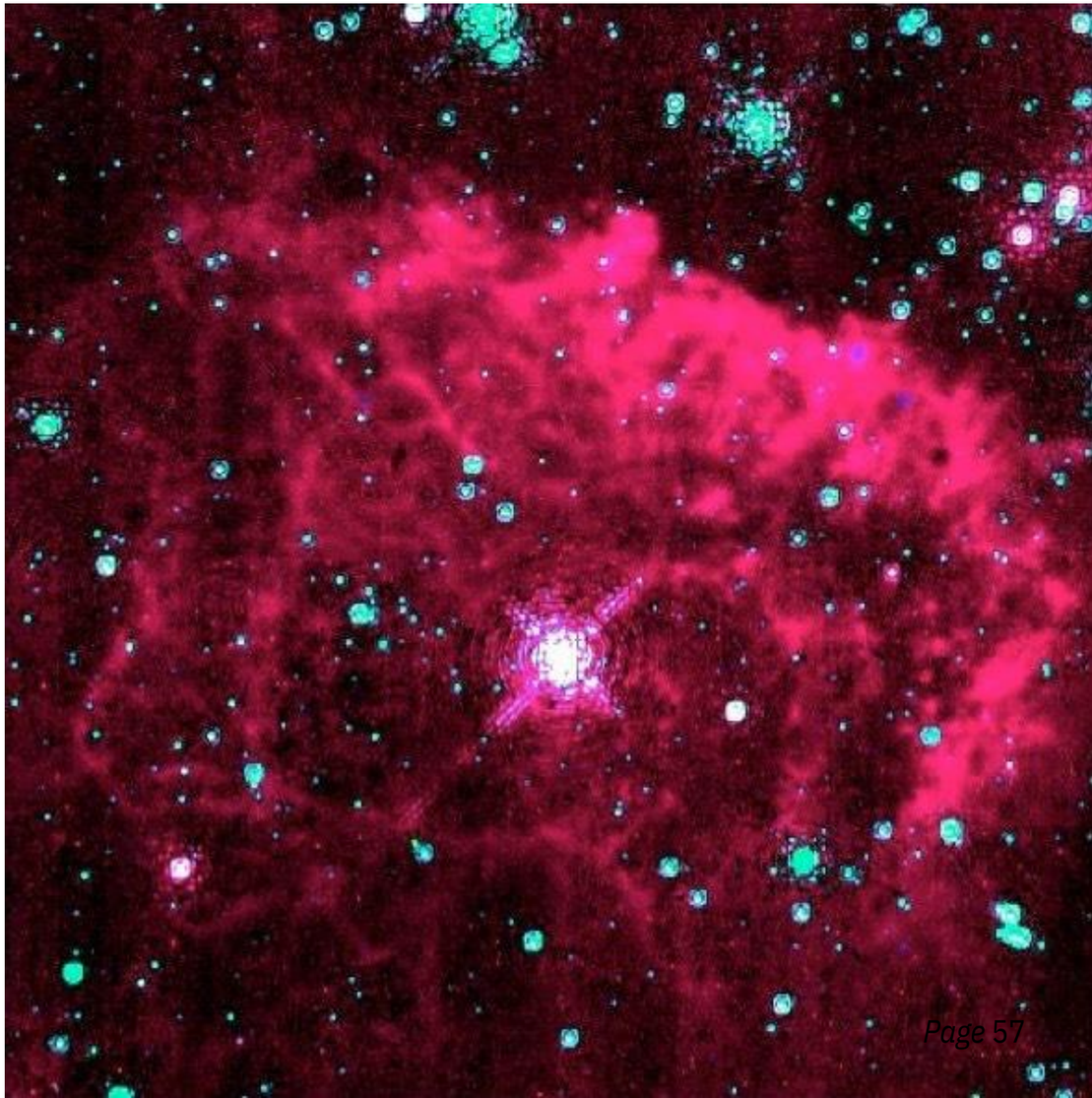
Credit: A. Heger, `https://2sn.org/stellarevolution/explain.gif`

Mass loss

One major complication in understanding the evolution of massive stars is that they lose large amounts of mass at many stages of their life. Hot stars emit a continuous outflow of matter from their surfaces as a *stellar wind*.

Our own Sun has a solar wind which reaches speeds of 400–700 km/s with a mass loss rate of about 10^{-14} solar masses per year.
Over a ten billion year lifespan, at this rate the Sun will lose about 0.01% of its mass to the solar wind.

By contrast, the winds from hot stars can be a billion times stronger, losing up to 10^{-5} solar masses per year at speeds of up to 3000 km/s. This means that even during the much shorter lives of the stars (a few million years), they can lose on the order of half or more of their mass.



The "Pistol Nebula" and its central star, which may have weighed up to 200 times the mass of the Sun before shedding much of its mass in violent eruptions.

The Bubble Nebula, NGC 7635, is being pushed out by the stellar wind of massive central star BD+602522, which has a mass about 40 times the mass of the Sun. The bubble is about 3 parsecs (10 ly) across.

How do the most massive stars lose mass?

What causes this mass loss, considering that these stars — unlike the Sun, where a very small mass loss occurs due to convection in the photosphere — have radiative envelopes?

Let us consider the pressure balance

$$\frac{dP}{dR} = -\frac{GM\rho}{R^2}$$

and the radiation pressure, which in the absence of convection satisfies

$$\frac{dP_{rad}}{dR} = -\frac{kL\rho}{4\pi cR^2}$$

you obtain if you substitute $$\frac{dT}{dr} = -\frac{3}{4ac}\frac{k\rho}{T^3}\frac{L_r}{4\pi r^2}$$

In the derivative of $$P_{rad} = \frac{1}{3}aT_c^4$$

So, you can obtain $$\frac{dP_{rad}}{dP} = \frac{kL}{4\pi GMc}$$

Therefore: $$\frac{dP_{rad}}{dP} < 1$$ being $P=P_{rad}+P_{gas}$

So, we will have on the surface of the star

$$L_s < \frac{4\pi cGM_s}{k_s}$$ = Eddington luminosity

This limit can be reached $P_{rad} >> P_{gas}$

In massive stars, P_{rad} is significant, and their luminosity is a non-negligible fraction of the Eddington luminosity. If we assume a mass-loss rate $\frac{dM_s}{dt}$,this mass must leave the star of mass $M_s$with at least the escape velocity. Therefore the energy loss will be:

$$L_{\dot{M}} = \frac{GM_s\dot{M}_s}{R_s}$$

If we define the stellar luminosity L_s as a fraction λ of the Eddington luminosity, we have:

$$\frac{L_{\dot{M}}}{L_s} = \frac{k_s\dot{M}_s}{4\pi\lambda cR_s}$$

where the ratio λ is observationally < 1, but not by much.

LINE DRIVEN winds

In the outer atmosphere of the star, atoms are found in the appropriate ionization and excitation state to absorb radiation at a frequency close to the star's peak emissivity.

Statistically, there will be a net outward gain in momentum. The atoms responsible for absorbing the radiation must then transfer their momentum to the surrounding material through collisions, so that the matter as a whole acquires sufficient momentum to escape from the star.

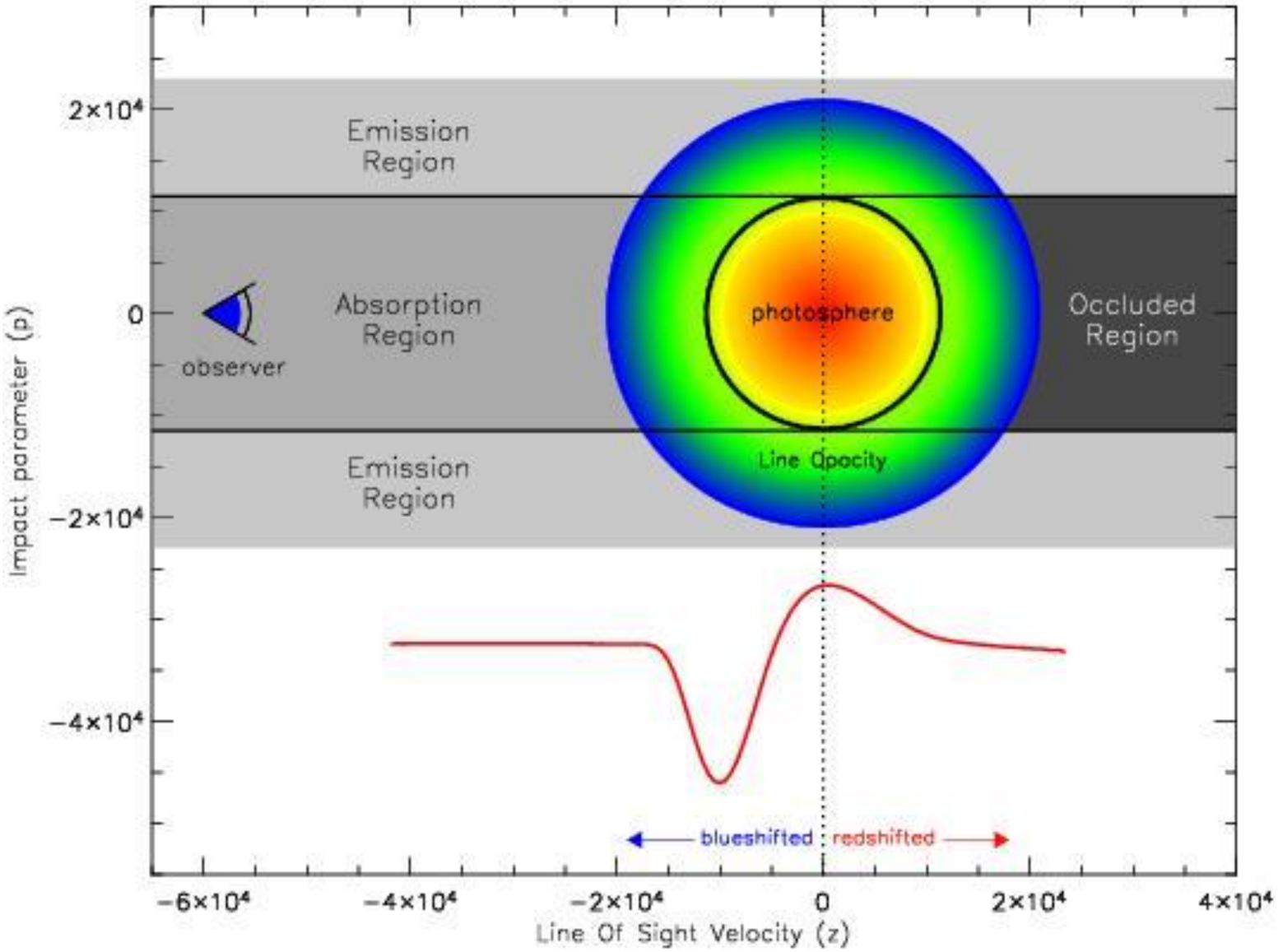
In the atmospheres of massive stars, H and He are fully ionized, and therefore the atoms that absorb the radiation are metals.

Measurements of mass loss

Massive stars often show P-Cygni profiles in their spectral lines.

These profiles consist of an absorption component (P) and an emission component (Cygni).

The absorption component indicates material moving toward the observer, while the emission component indicates material moving away.

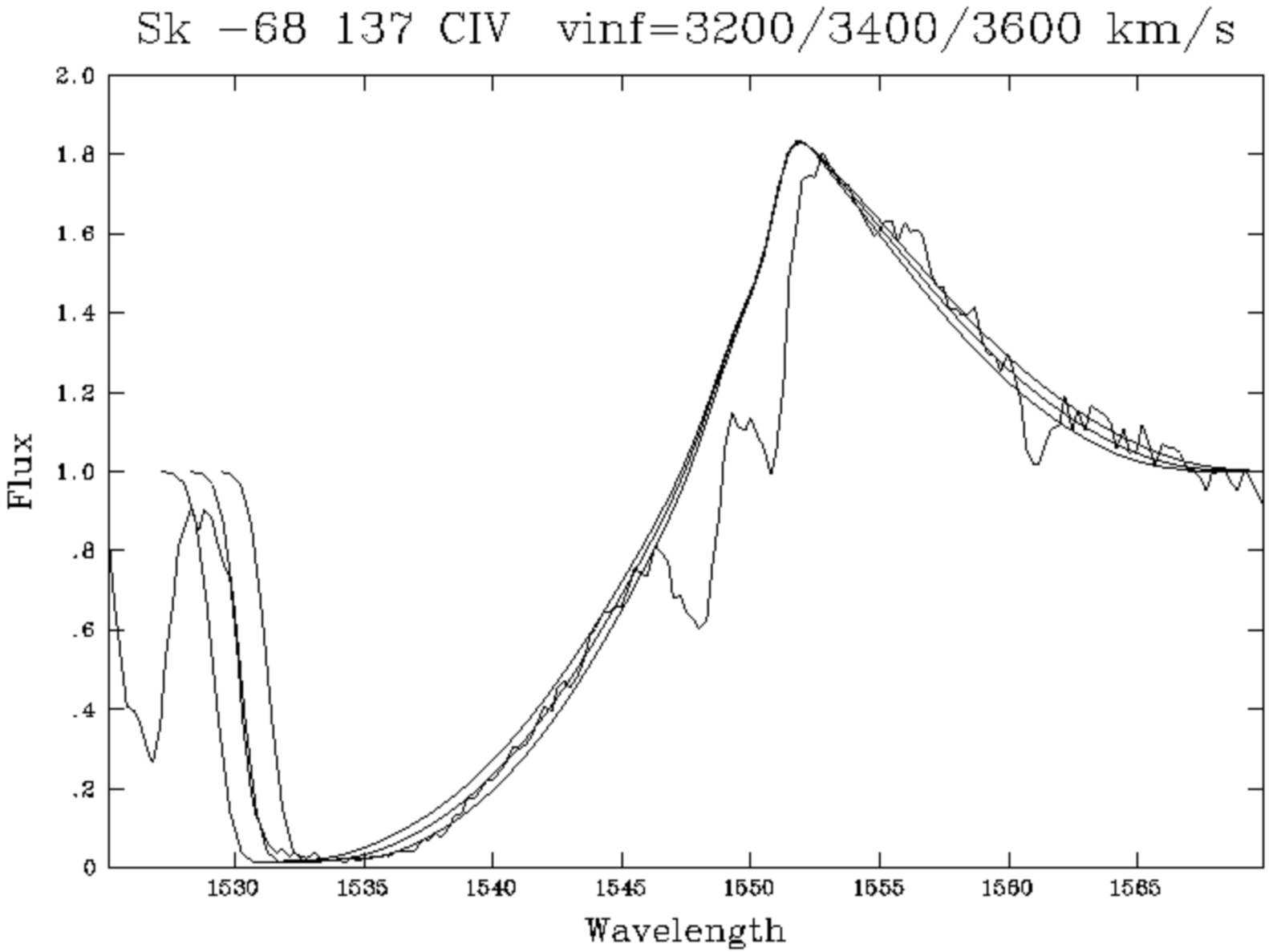


Measurements of mass loss

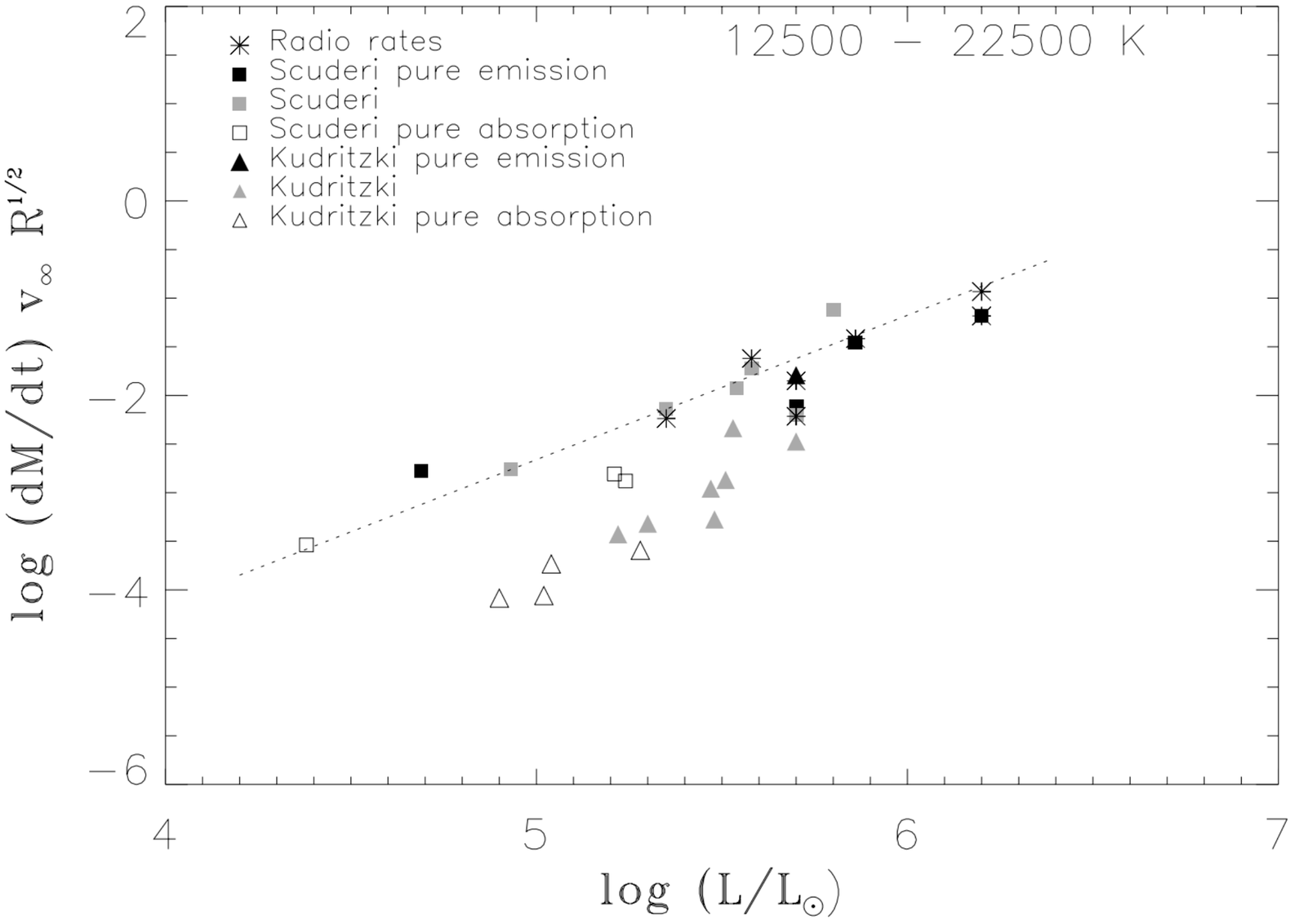
Example of a velocity measurement from Kudritzki (1998) for an O3 star in the LMC.

$$\dot{M} = 4\pi\rho(R)v(R)R^2 =$$

If we know the density and the radius, we can obtain an estimate, but it is often not straightforward to convert from velocity to mass loss.



Observational data



EVOLVED MASSIVE STARS II: HOT TYPES

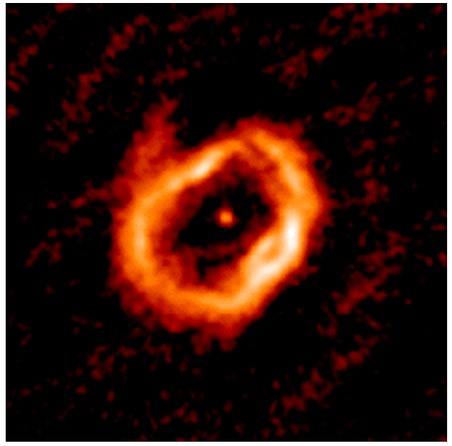
Wolf-Rayet star (WR): He-burning massive stars with very strong emission lines of He in their spectra, caused by very strong winds. WN (WR with N) and WC (with C lines) are exposed He cores of massive stars that have lost their H envelopes through prior mass loss. *Examples: γ^2 Vel, EZ CMa.*

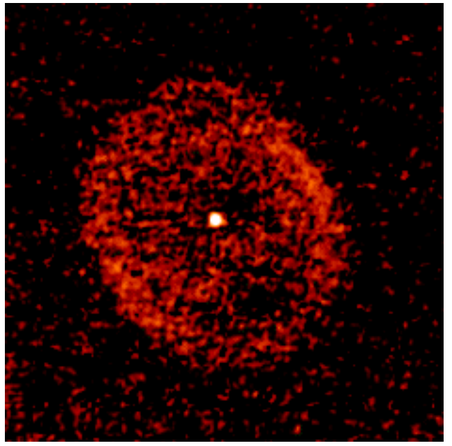
Luminous blue variable (LBV): A group of evolved massive stars that exhibit eruptive mass loss or irregular variability. A union of various subtypes, including giant eruptions (η Car variables), S Dor variables, α Cyg variables, P Cygni stars, Hubble-Sandage variables, etc. Most have strong winds and strong emission-line spectra. Candidate LBVs are stars that have similar spectra and/or dust shells, but have not yet been seen to exhibit variability. *Examples: η Car, P Cygni, AG Car, S Dor, HR Car.*

Blue supergiant (BSG): Post-MS massive stars with B spectral types. The relative number of BSGs in observed HR diagrams of stellar populations is not well understood. *Examples: Sk-69 202, Sher 25, SBW1.*

Be and B[e] stars: B-type stars with strong and usually time-variable emission lines, often showing evidence for disk-like circumstellar material. Be stars are rapid rotators, possibly resulting from increased angular momentum through mass accretion in binaries. The B[e] stars have strong forbidden line emission and IR excess from dust that are thought to arise in a circumstellar disk or torus. Some are high-luminosity evolved supergiants similar to LBVs. *Examples: γ Cas (Be), R4 in the SMC (B[e]).*

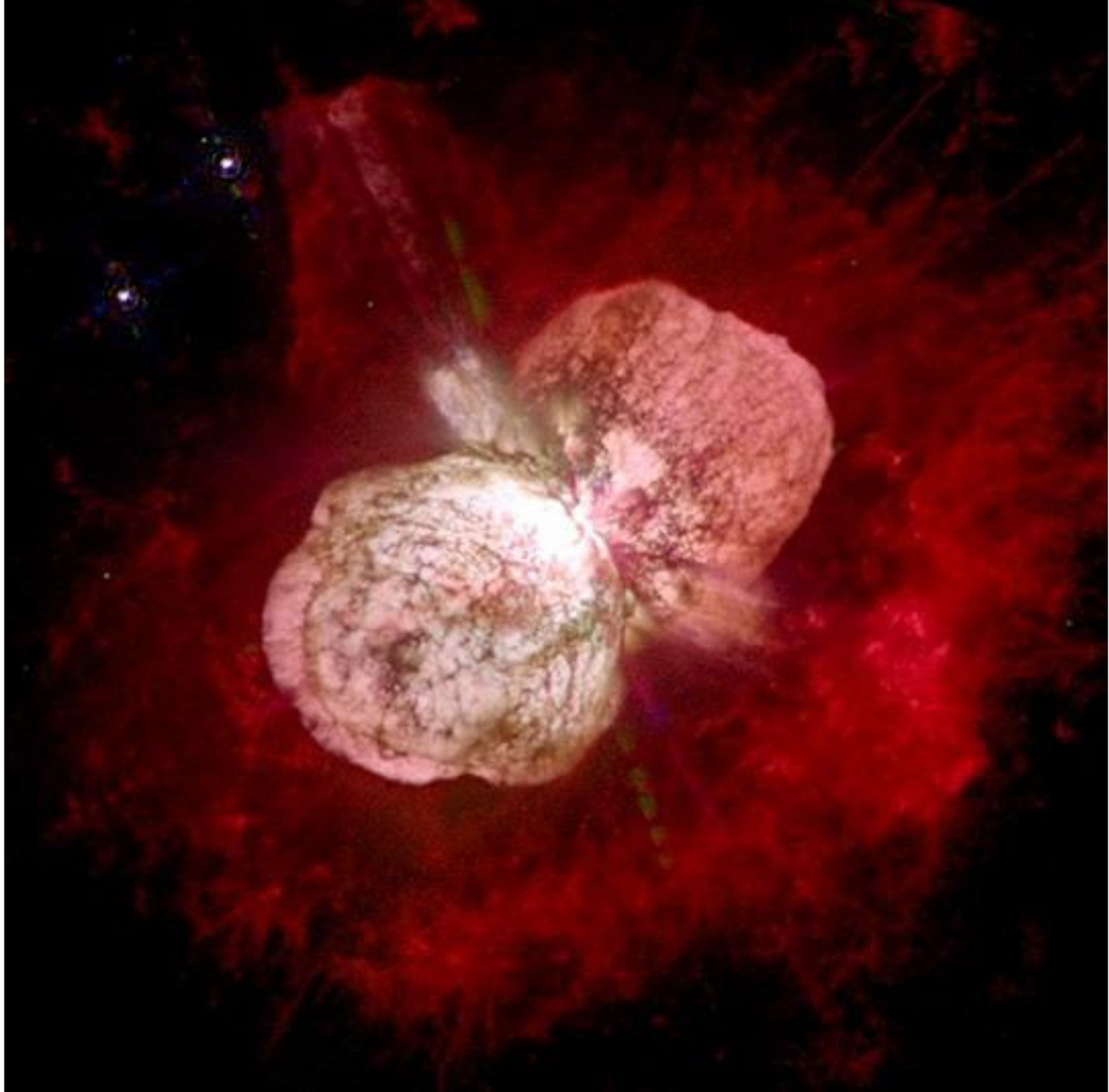
Very massive stars, with initial masses greater than 40 solar masses, become *luminous blue variables* after leaving the main sequence. During this phase, which lasts for perhaps 40,000 years, the stars are highly variable and losing mass through strong winds. At minimum brightness they appear as blue B-type supergiants; during outburst they are much redder. Every few centuries they have sudden giant eruptions, ejecting large amounts of mass.

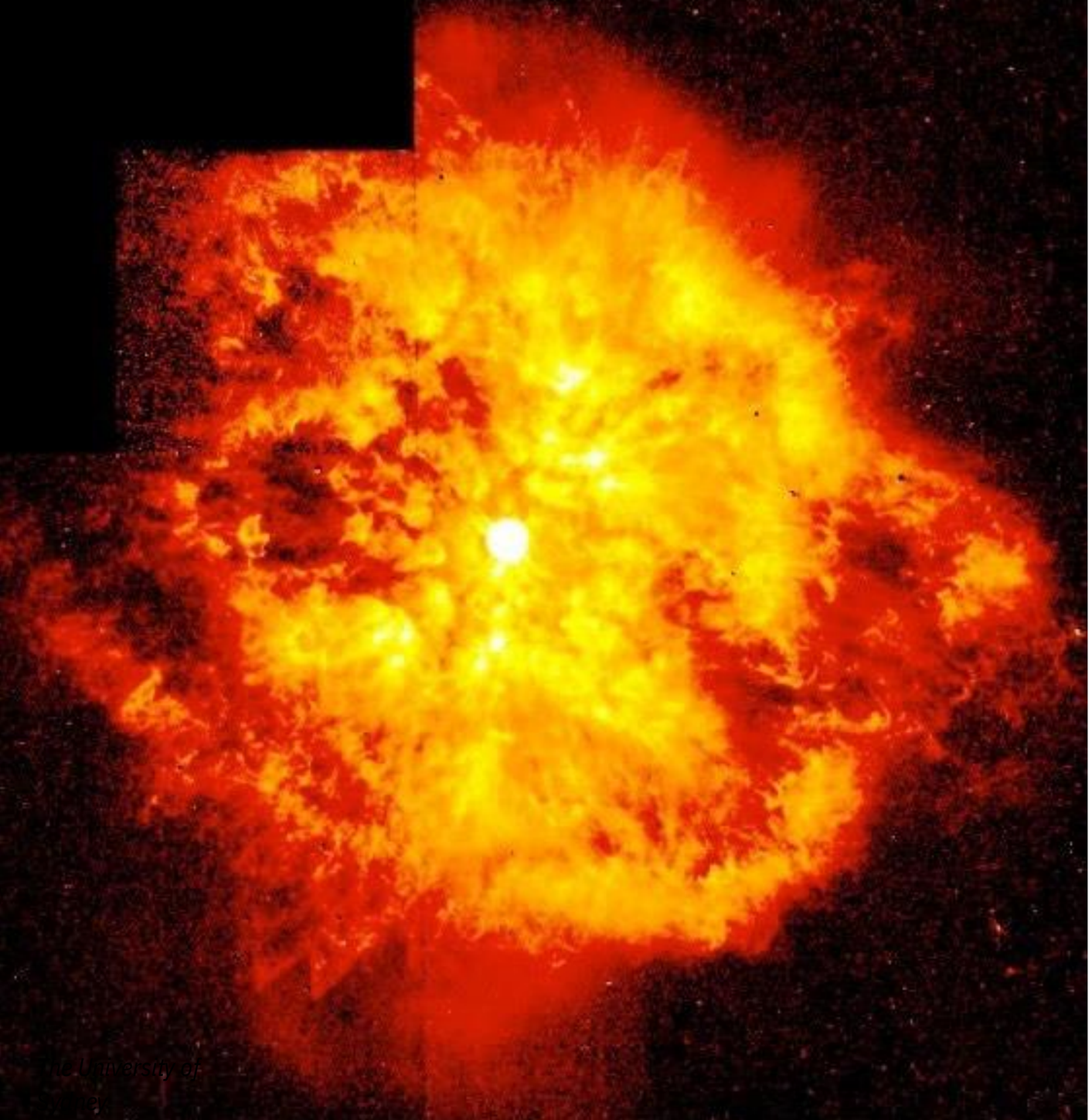




Radio images of two luminous blue variables, AG Carinae and Henize 3–519, showing rings of emission from mass lost during major eruptions.

Eta Carinae is the most famous luminous blue variable. During the 1840s it brightened by 4 magnitudes, becoming one of the brightest stars in the sky. Hubble images show two huge bubbles of gas, remnant of the expulsion of about a solar mass of material.



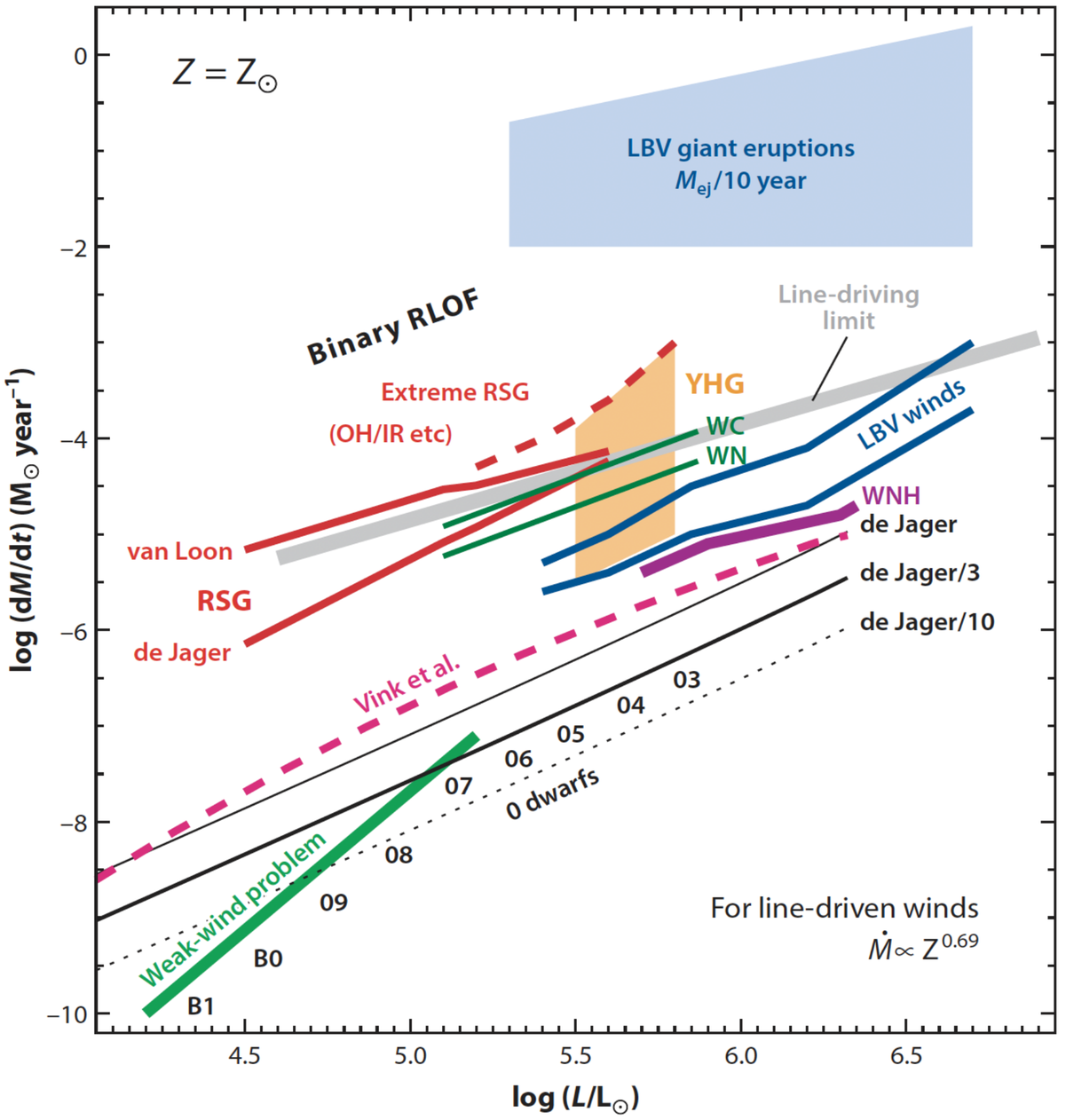


Wolf-Rayet stars represent the most extreme stage of mass loss in the life of a massive star: the star has lost so much mass that they are actually exposing the underlying layers which have already undergone nuclear fusion.

HST image of the Wolf-Rayet star WR124, showing the star surrounded by hot clumps of gas being ejected at high speed.

A number of different prescriptions for wind mass loss used in models, as well as typical observed ranges of mass-loss rates for a number of different types of stars. For O-type stars, the theoretical rates from the prescription by Vink et al. (2001) are shown, along with "standard" observational rates using the prescription by de Jager et al. (1988), as well as these same prescriptions divided by factors of 3 and 10 for comparison.

The green line labeled "weak-wind problem" refers to lower mass-loss rates for late O-type and early B-type MS stars. Rates for nitrogen-sequenceWolf-Rayet (WN) and carbon-sequenceWolf-Rayet (WC) stars are from Crowther (2007). Red supergiant (RSG) mass-loss prescriptions are from de Jager et al. (1988) and van Loon et al. (2005), as indicated. For yellow hypergiants (YHGs), see de Jager (1998). For $\dot{M}$ corresponding to normal winds of luminous blue variables (LBVs), values were compiled from a number of studies (Hillier et al. 2001, Vink & de Koter 2002, Smith et al. 2004, Groh et al. 2009). For LBV eruptions, the "rates" shown are calculated from total masses observed in LBV circumstellar shells (Smith & Owocki 2006) divided by a nominal eruption duration of 10 years. For "binary RLOF," an order-of-magnitude value for the strongest mass-transfer rates expected in brief RLOF (Roche-lobe overflow) phases is noted, although the mass-transfer or mass-loss rate can be much less for slow mass transfer or possibly more for dynamical common-envelope ejection events; see references in the text, especially the review by Langer (2012).



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19 Novembre 2025 alle ore 17:00.

Paolo Molaro **(Osservatorio Astronomico di Trieste)**

Massimiliano e la nascita dell'Osservatorio Astronomico di Trieste

https://www.ifpu.it/miramare25/

Mass-loss prescriptions for hot stars (of the early spectral types)

For massive stars (the ones we have discussed so far), we typically have, as seen, power-law relations of the form

$$\dot{M} = L^{a}$$

or, more generally,

$$\dot{M} = F(M, L, R, Z)$$

and it is generally linked to radiation pressure.

Prescription of stellar winds for evolved stars (giants and supergiants)

What causes mass loss in cool stars?

Radiation pressure on dust grains and molecules

Solar-type mass loss, i.e., driven by the pressure of hot gas

Mass loss caused by shock waves

The mass loss can be parameterized using the **Reimers parameter** η (typically $0.3-3$)

$$\dot{M} = \frac{\eta \cdot 10^{-13} L}{gR}$$

or alternatively using the formulation by **Fusi Pecci and Renzini**, with $\eta_{FR} = 8 \times 10^{-4}$and the **acoustic luminosity**

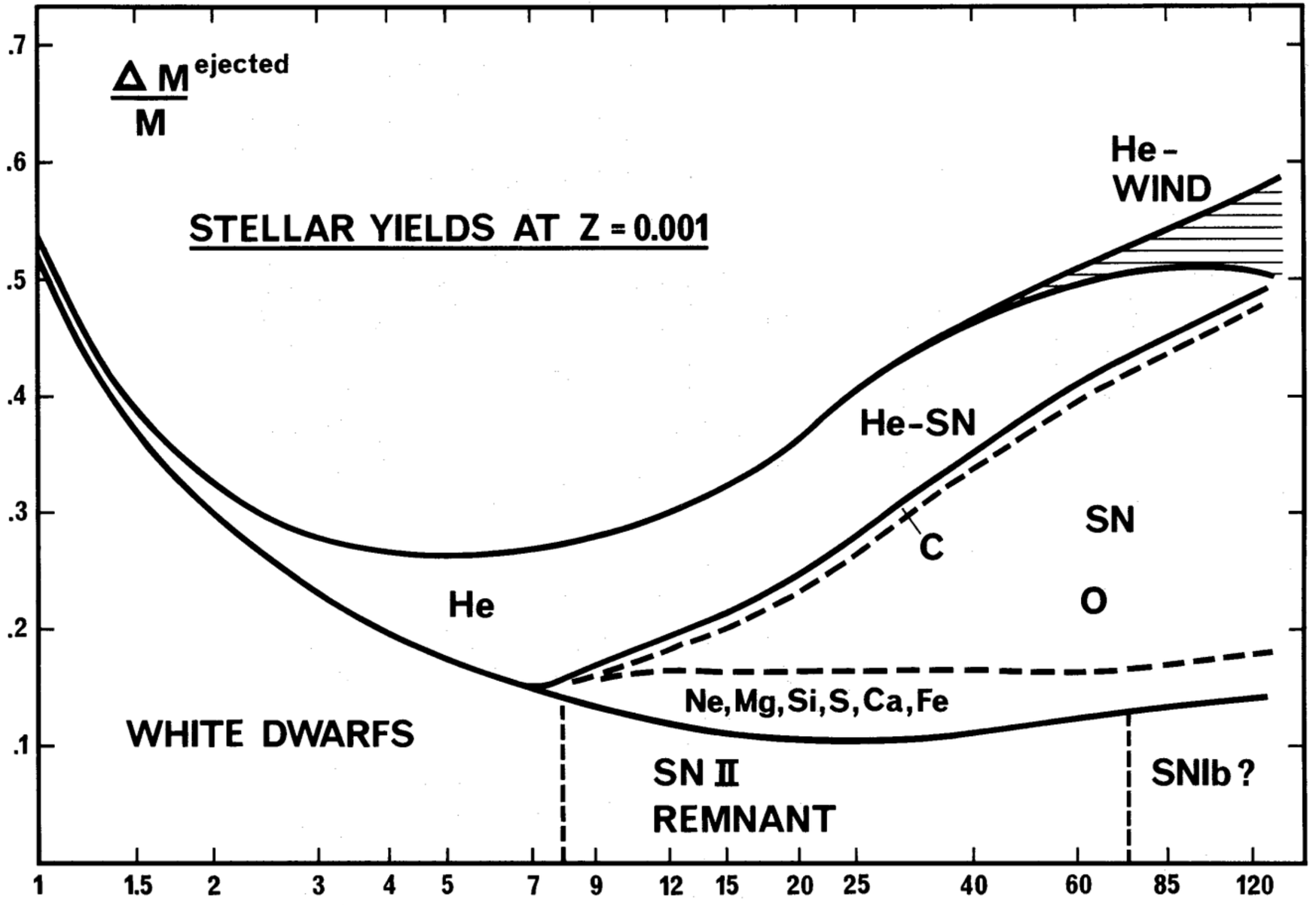
$$\dot{M} = \frac{\eta_{FR} L_a}{gR}$$

Mass loss has a strong implications for the evolution of the star. Reducing the mass of the star reduces the pressure and temperature in the interior, which can reduce the mass of the core.

The amount of mass lost in these stages of a star's life essentially determines what its ultimate fate is and we note here that most of the uncertainty in what mass of star leads to what kind remnant is because of the uncertainty in how much mass is lost by the time the star ends its life.

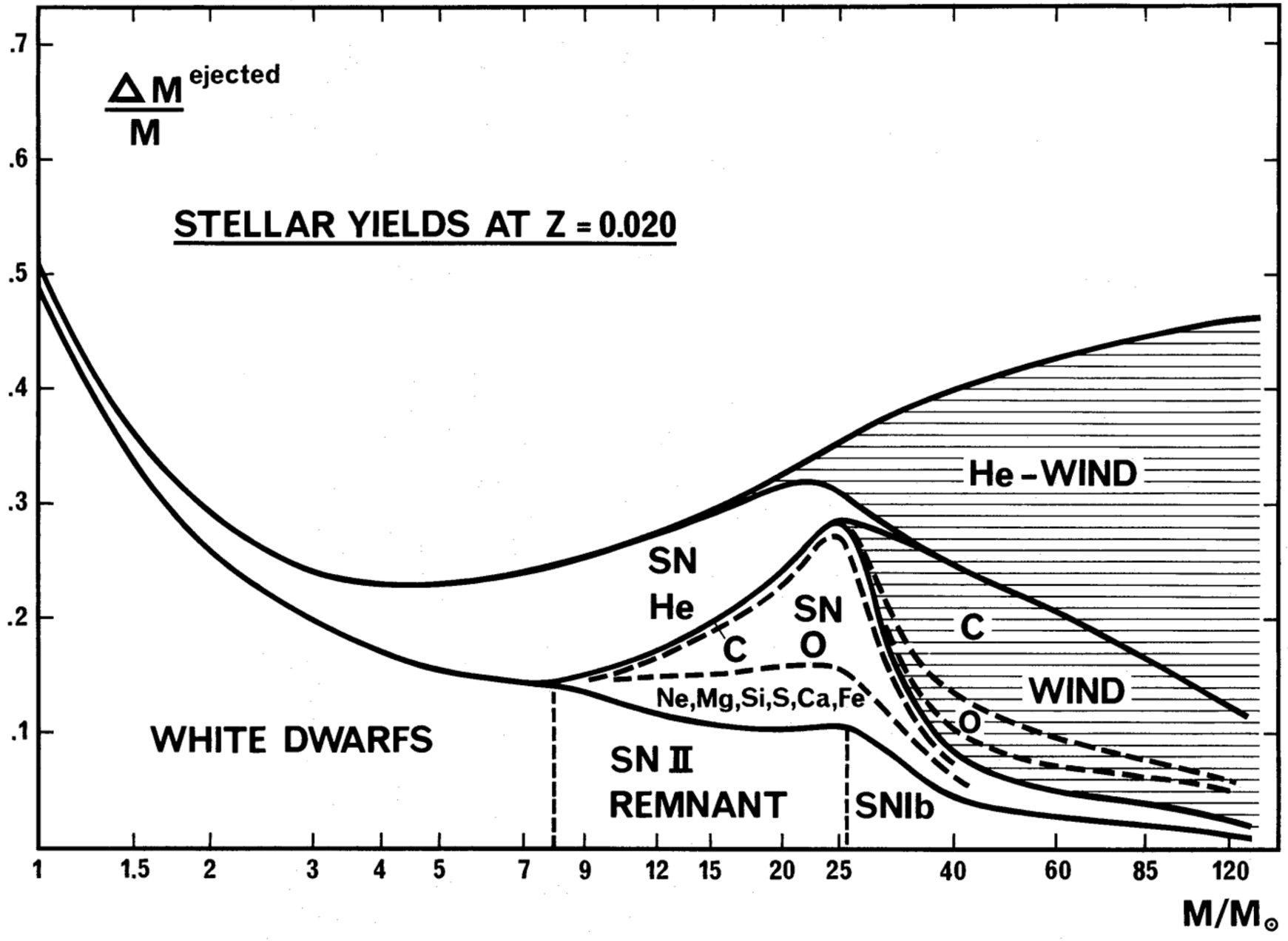
Maeder 1992

At low metallicity,
(almost)
no mass loss

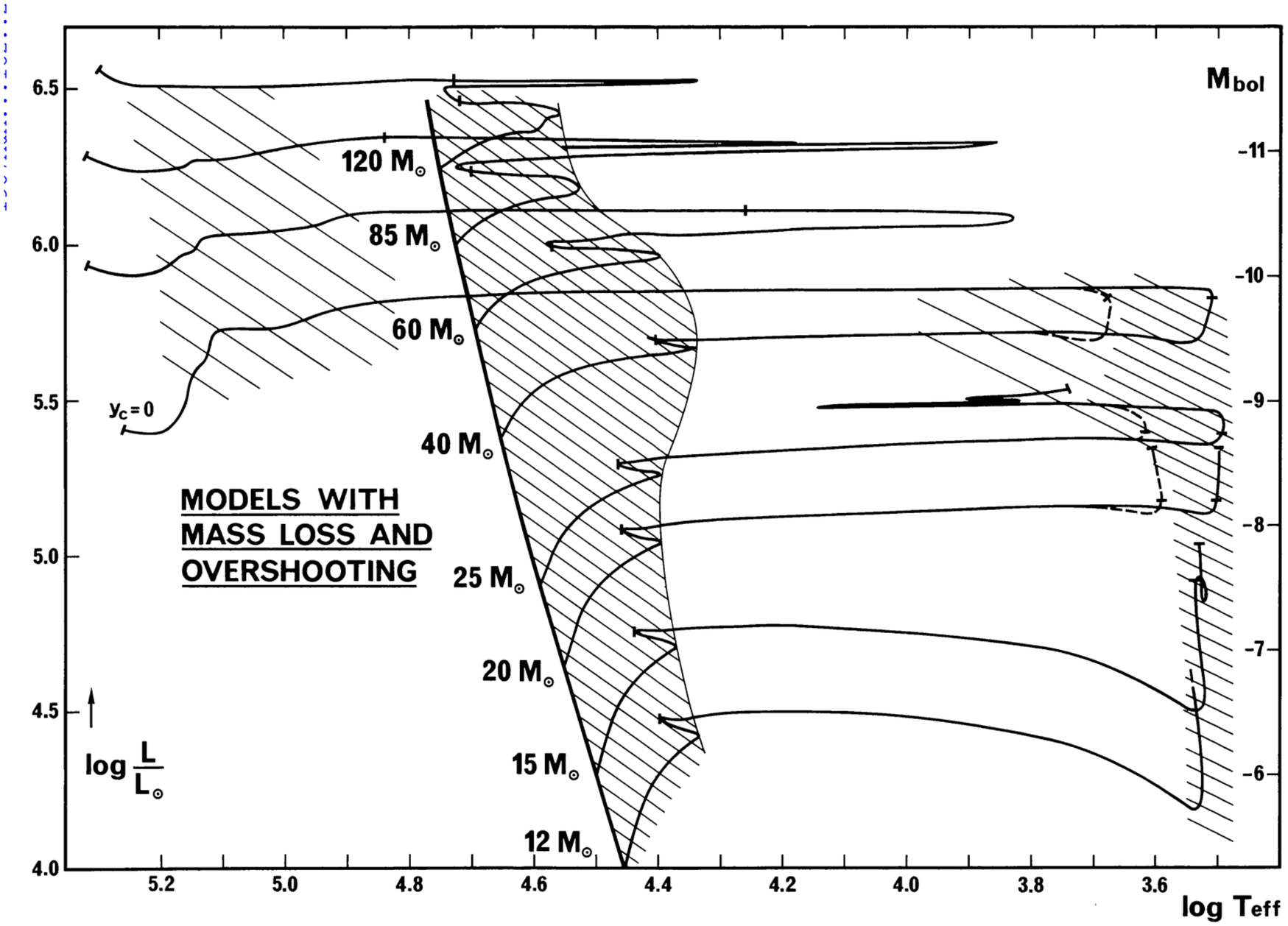


Maeder 1992

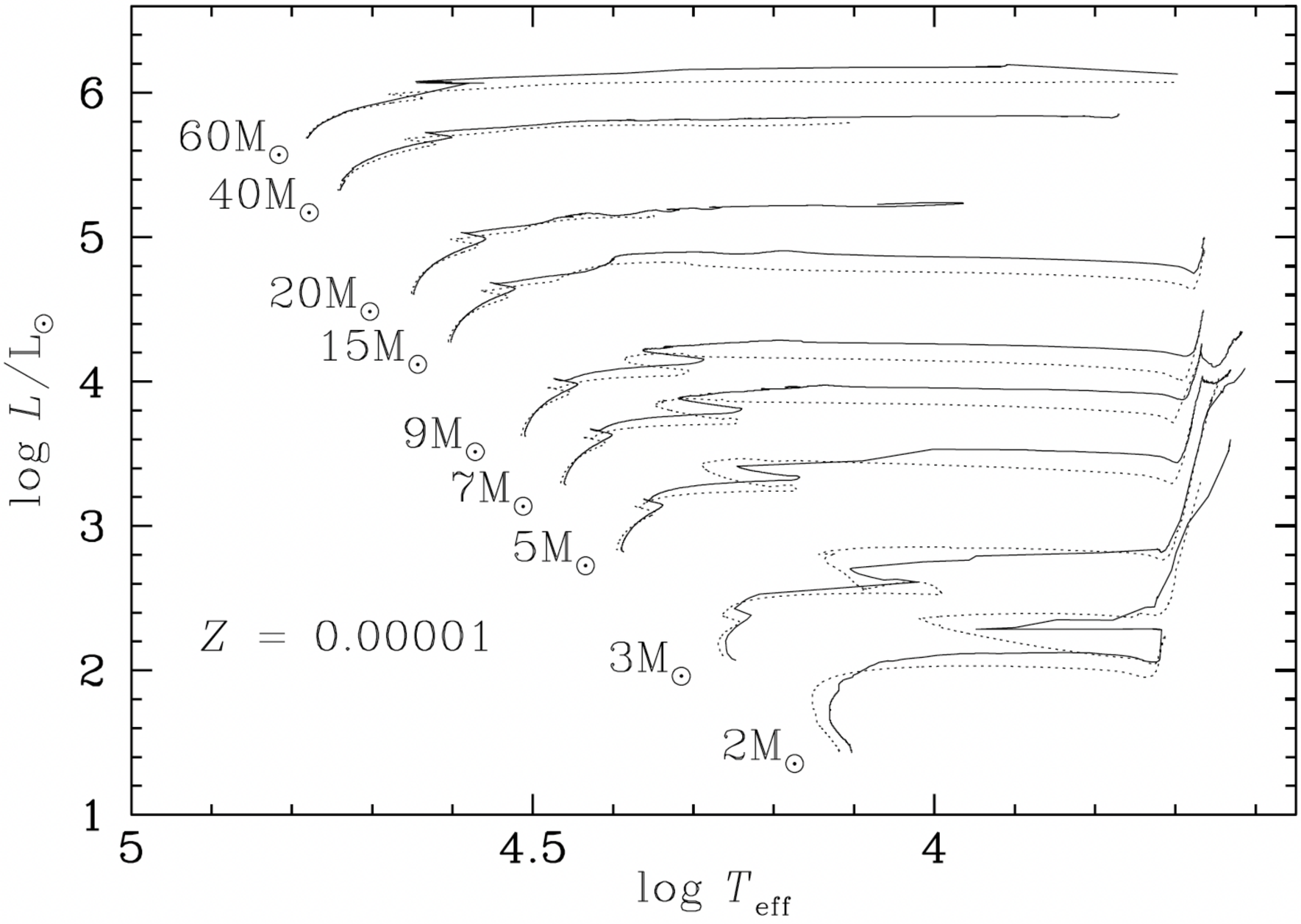
At solar metallicity the impact is quite visible in this plot, and implies also different chemical enrichment!



Maeder


120 M⊙
85 M⊙
60 M⊙
40 M⊙
25 M⊙
20 M⊙
15 M⊙
12 M⊙
y_c = 0
MODELS WITH MASS LOSS AND OVERSHOOTING
log L/L⊙
M_bol
log Teff
6.5
6.0
5.5
5.0
4.5
4.0
-11
-10
-9
-8
-7
-6
5.2
5.0
4.8
4.6
4.4
4.2
4.0
3.8
3.6

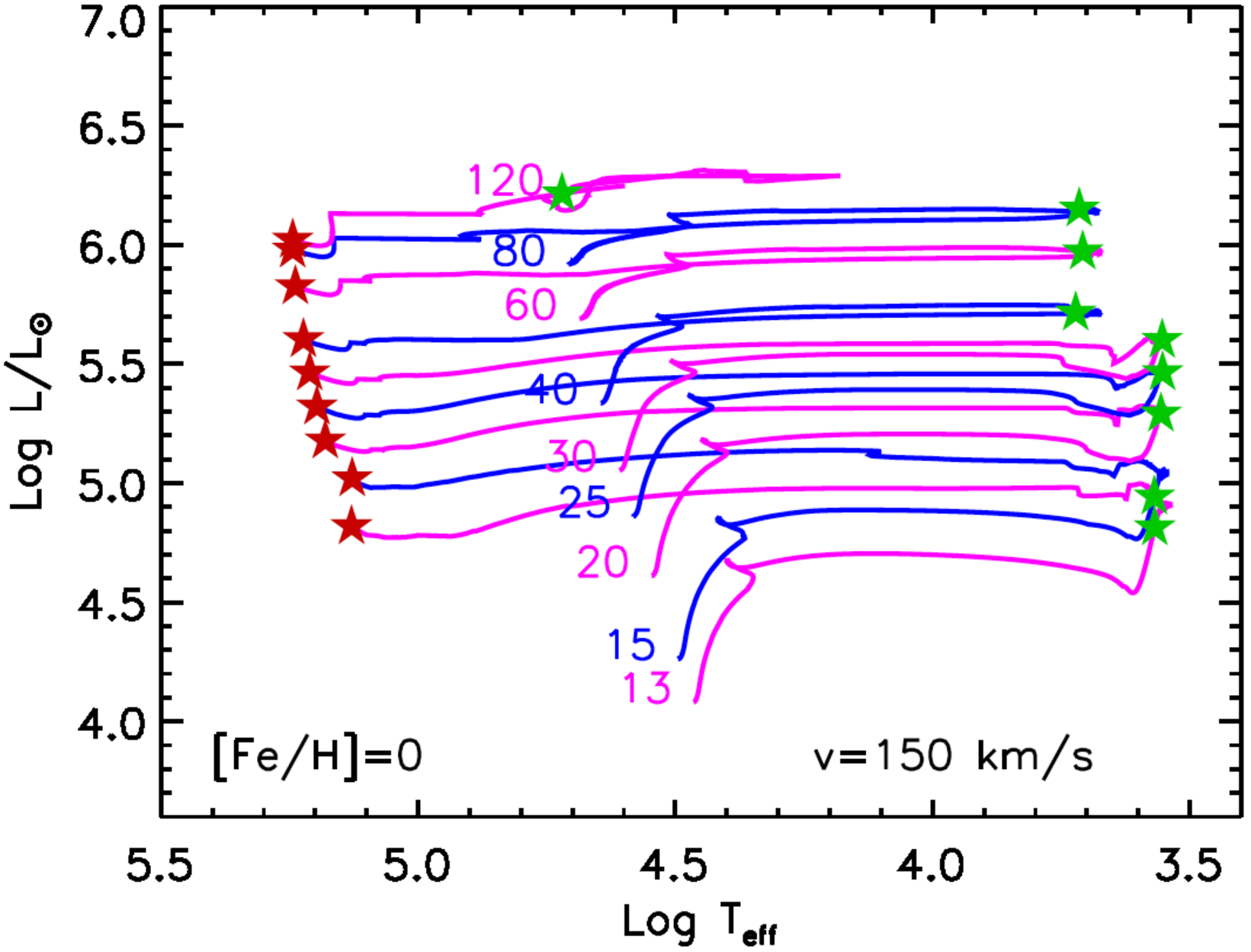
Rotazione Meynet & Maeder 2002

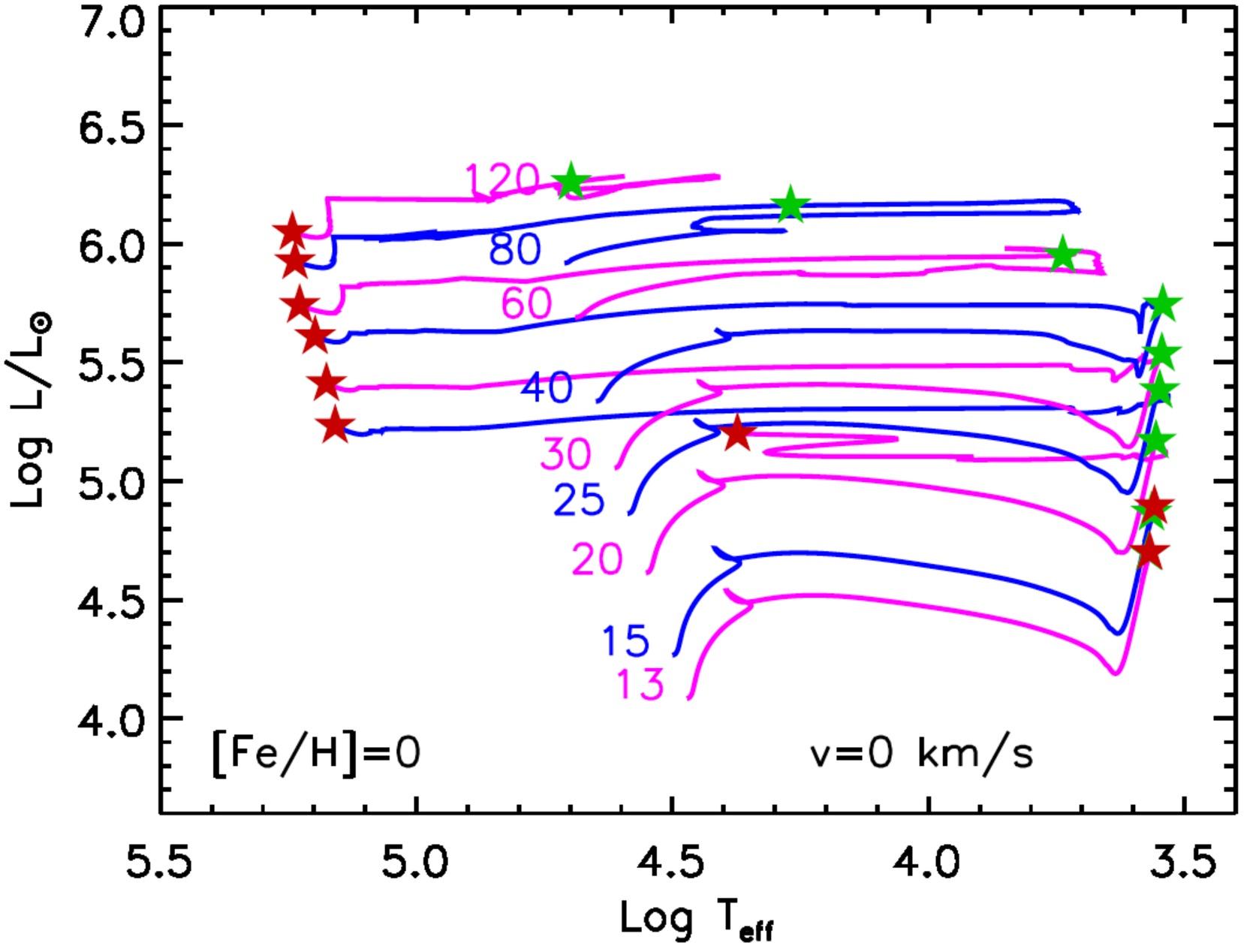


Rotazione Limongi 2018

It is crucial for the star to explode as a red supergiant;

SN 1987A exploded as a **blue supergiant** (whereas non-rotating models predict an explosion as a red supergiant at ~20 M⊙).




7.0
6.5
6.0
5.5
5.0
4.5
4.0
Log L/L⊙
120
80
60
40
30
25
20
15
13
[Fe/H]=0
v=0 km/s
5.5
5.0
4.5
4.0
3.5
Log T_eff

Rotation tends to:

- Lower the surface gravity due to centrifugal force and angular momentum transport, which shifts the evolutionary track toward redder colors.
- Induce mixing that increases the size of the region where hydrogen has been burned, moving the track to higher luminosities and lower temperatures.

At the same time, the radiative envelope is enriched in helium, which reduces opacity and makes the track brighter and bluer.
Rotating massive stars produce primary nitrogen thanks to rotation: the star must have a helium-burning core and a dense, long-lived hydrogen-burning shell. The diffusion and transport of fresh carbon from the core to the shell allow the production of primary nitrogen from this fresh carbon through hydrogen burning via the CNO cycle.

The fate of the envelope is to be ejected in a powerful explosion:

SN II, SN Ib, or SN Ic.

Why does the explosion occur? We are getting there (see the work of Janka, MPA). In any case, neutrinos are fundamental, as well as the effects of asymmetry (1D simulations fail to produce explosions).

The fate of the envelope is to be ejected in a powerful explosion:

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During the explosion, all the metals produced during the star's lifetime are ejected into the interstellar medium, giving rise to the chemical evolution of the galaxy.

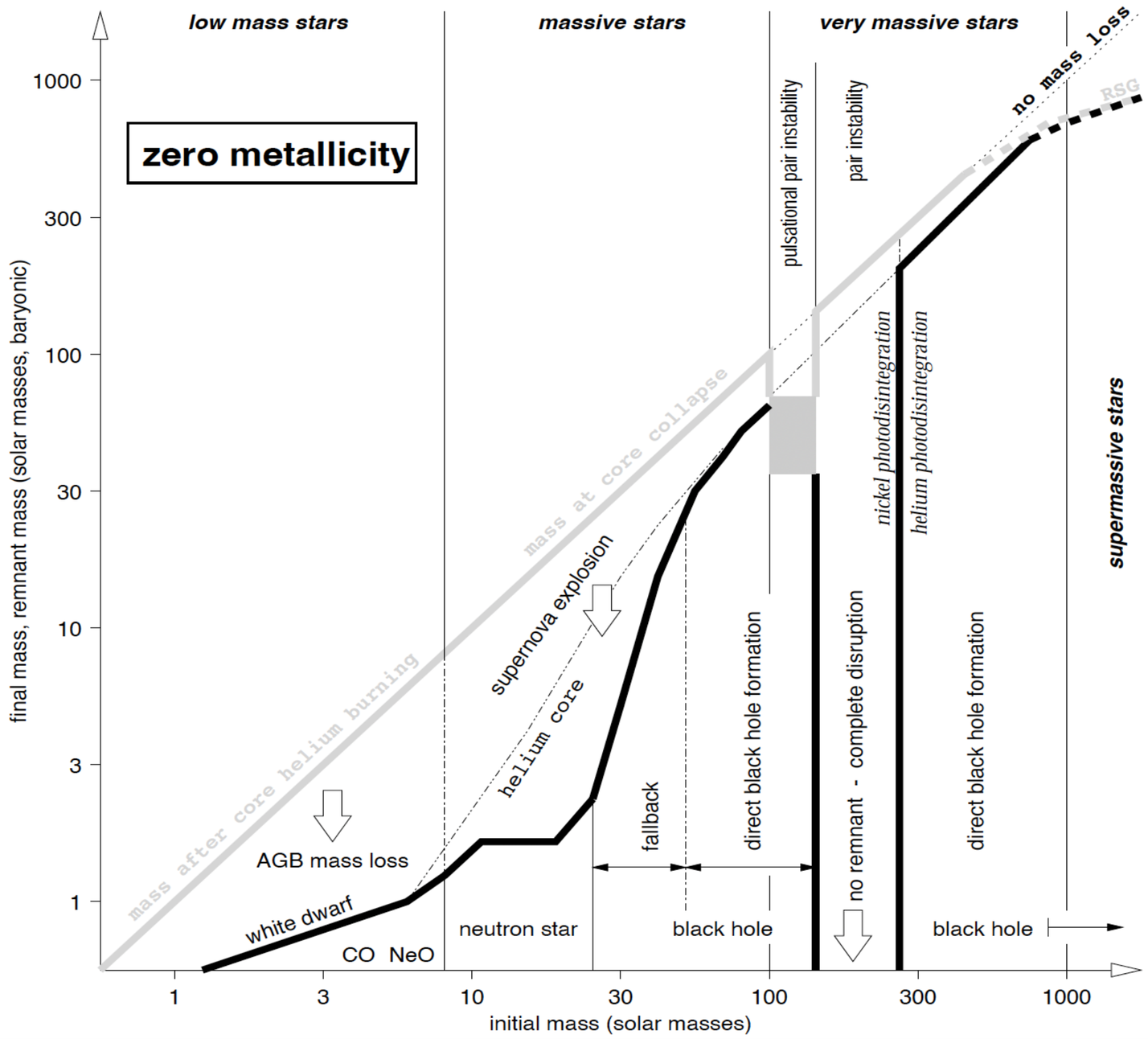
[Submitted on 12 Jun 2012]

Explosion Mechanisms of Core-Collapse Supernovae

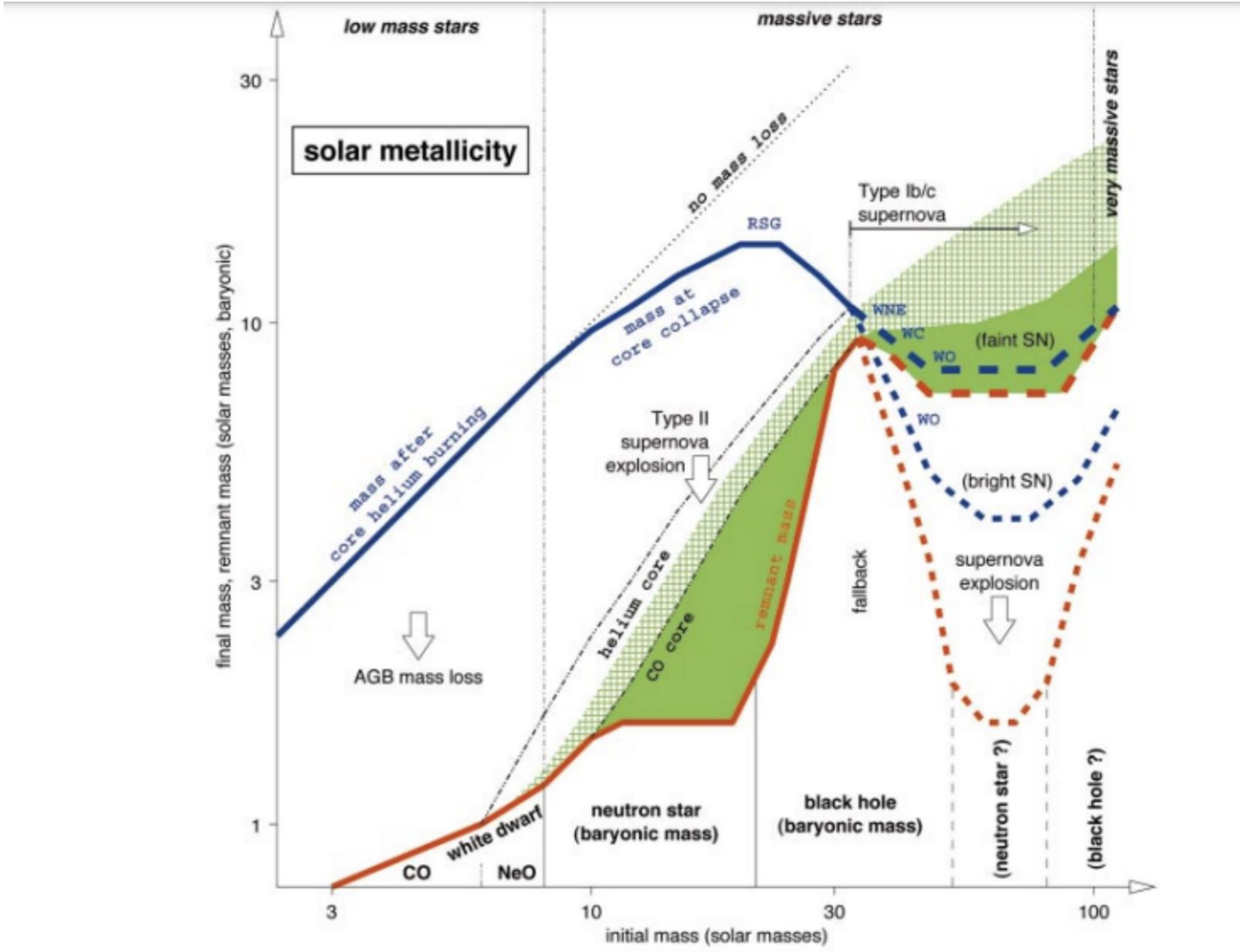
H.-Thomas Janka (Max Planck Institute for Astrophysics, Garching)

Supernova theory, numerical and analytic, has made remarkable progress in the past decade. This progress was made possible by more sophisticated simulation tools, especially for neutrino transport, improved microphysics, and deeper insights into the role of hydrodynamic instabilities. Violent, large-scale nonradial mass motions are generic in supernova cores. The neutrino-heating mechanism, aided by nonradial flows, drives explosions, albeit low-energy ones, of ONeMg-core and some Fe-core progenitors. The characteristics of the neutrino emission from new-born neutron stars were revised, new features of the gravitational-wave signals were discovered, our notion of supernova nucleosynthesis was shattered, and our understanding of pulsar kicks and explosion asymmetries was significantly improved. But simulations also suggest that neutrino-powered explosions might not explain the most energetic supernovae and hypernovae, which seem to demand magnetorotational driving. Now that modeling is being advanced from two to three dimensions, more realism, new perspectives, and hopefully answers to long-standing questions are coming into reach.

Fate of a star



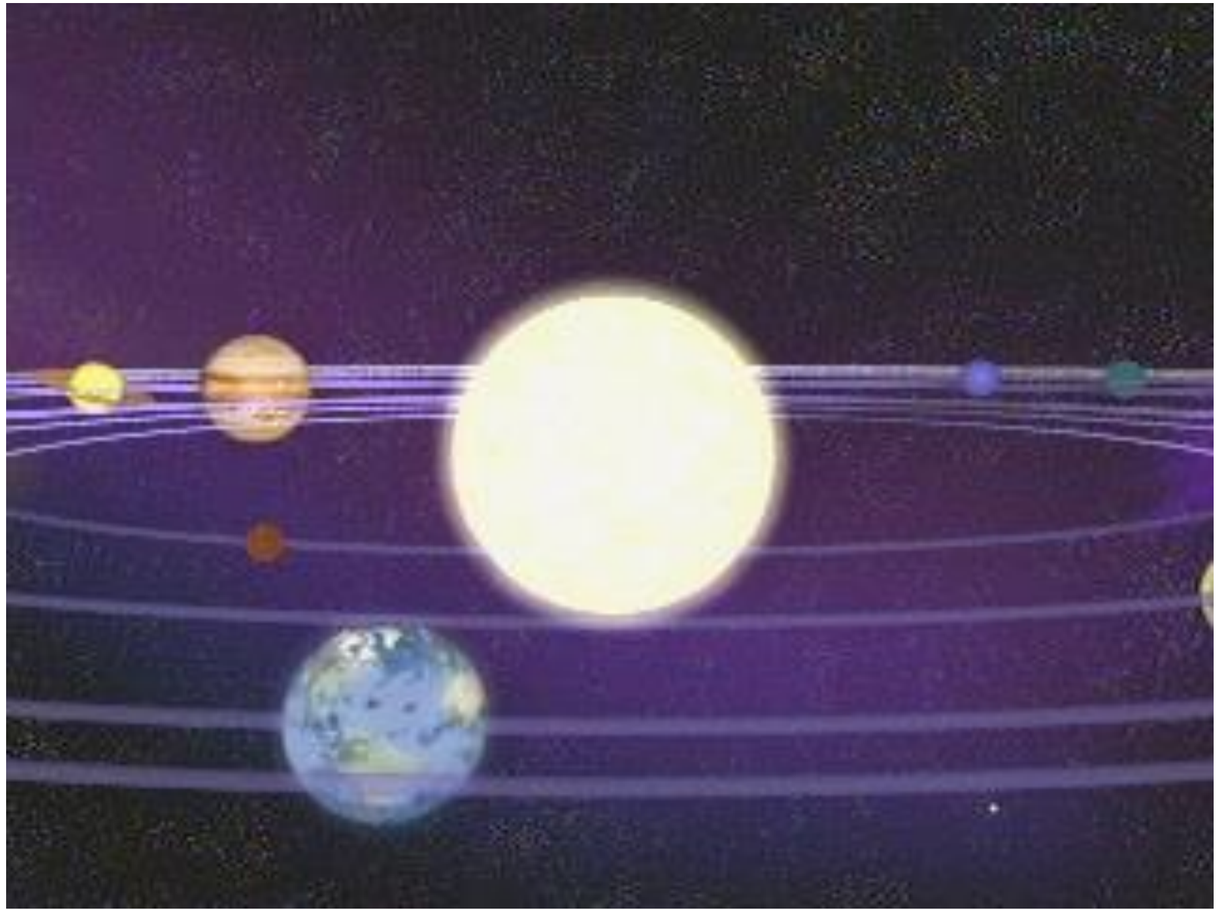
Fate of a star



Credit: Woosley, Heger & Weaver (2002)

The final fate of the Earth

Mass loss is also vital for determining the final fate of the Earth. As the Sun ages, it will expand and become a red giant.



The final radius the Sun reaches is important. However, as the Sun loses mass, its gravitational pull on the planets weakens, and their orbits expand.

Will they expand enough to keep ahead of the expanding photosphere, or will they be engulfed?

Recent calculations* suggest that although the Sun's outer surface may not quite reach Earth, tidal interaction between the Earth and the giant Sun will drag the Earth inwards, to be engulfed by the Sun.

This will take place just before the Sun reaches the tip of the RGB, around 7.59 ± 0.05 Gyr from now.

In any case, the Earth is likely to become uninhabitable long before that point is reached.