# 3D Radiation MHD Simulations of Massive Star Envelopes


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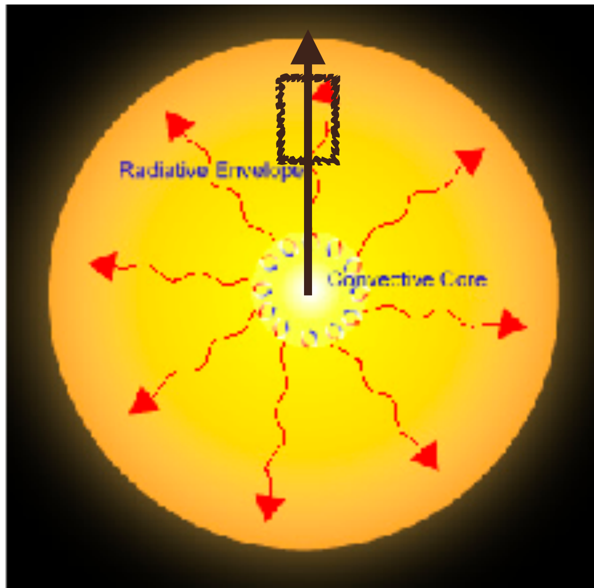
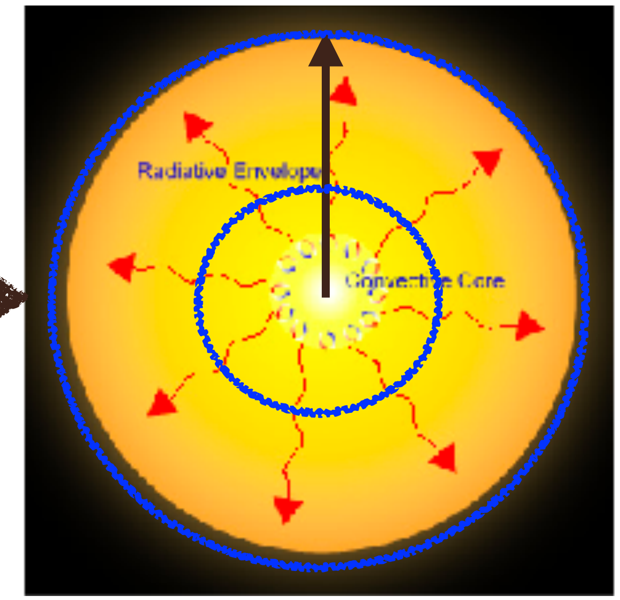
# Outline




- Numerical methods

- Local radiation MHD simulations of massive star envelopes

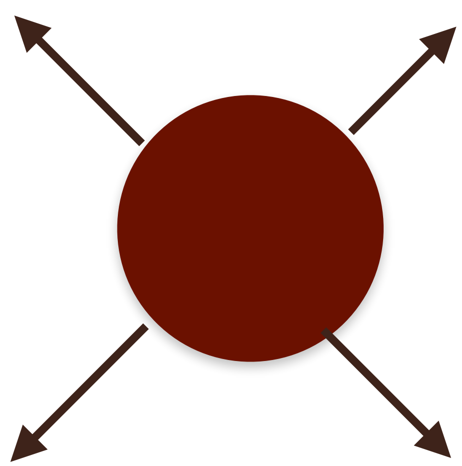
- Global simulations of massive star envelopes

# 1D Stellar Evolution Studies

Paxton et al. (2013)
Joss et al. (1973)

See Frank's, Matteo's talks

- Thermal equilibrium: how to transport the energy out



$$L = 4\pi r^2 F_r = 4\pi r^2 \left[F_{r,0} + v\left(E_r + P_r\right)\right]$$

Diffusive radiation flux

Advective flux

- Hydrostatic equilibrium

$$\frac{\mathrm{d}P_{\mathrm{gas}}}{\mathrm{d}r} = \left(\frac{\mathrm{d}P_{\mathrm{rad}}}{\mathrm{d}r}\right)\left[\frac{L_{\mathrm{Edd}}}{L_{\mathrm{rad}}} - 1\right].$$

$$a_r = \frac{\kappa F_{r,0}}{c}$$

$$\frac{dP_{\mathrm{rad}}}{dr} = -\rho a_r$$

# The Radiation MHD equations and Numerical Schemes

Jiang, Stone & Davis (2012)
Davis, Stone & Jiang (2012)
Jiang, Stone & Davis (2014)

Ideal MHD

$$\frac{\partial \rho}{\partial t} + \boldsymbol{\nabla} \cdot (\rho \boldsymbol{v}) = 0,$$

$$\frac{\partial (\rho \boldsymbol{v})}{\partial t} + \boldsymbol{\nabla} \cdot (\rho \boldsymbol{v}\boldsymbol{v} - \boldsymbol{B}\boldsymbol{B} + \mathsf{P}^*) = -\boldsymbol{S_r}(\boldsymbol{P}) - \rho \boldsymbol{\nabla}\phi,$$

$$\frac{\partial E}{\partial t} + \boldsymbol{\nabla} \cdot [(E + P^*)\boldsymbol{v} - \boldsymbol{B}(\boldsymbol{B} \cdot \boldsymbol{v})] = -cS_r(E) - \rho \boldsymbol{v} \cdot \boldsymbol{\nabla}\phi,$$

$$\frac{\partial \boldsymbol{B}}{\partial t} - \boldsymbol{\nabla} \times (\boldsymbol{v} \times \boldsymbol{B}) = 0. \qquad (1)$$

photon momentum

radiation energy

$$\frac{\partial I}{\partial t} + c\boldsymbol{n} \cdot \nabla I = S.$$

Radiative Transfer

$$S = c\rho\kappa_a \left( \frac{a_r T^4}{4\pi} - I_0 \right) + c\rho\kappa_s \left( J_0 - I_0 \right),$$

Absorption Scattering

# The Radiation MHD equations and Numerical Schemes

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$$\frac{\partial \boldsymbol{B}}{\partial t} - \nabla \times (\boldsymbol{v} \times \boldsymbol{B}) = 0. \quad (1)$$

photon momentum

radiation energy

Radiation moments

$$\frac{\partial E_r}{\partial t} + \nabla \cdot \boldsymbol{F_r} = cS_r(E),$$

$$\frac{1}{c^2}\frac{\partial \boldsymbol{F_r}}{\partial t} + \nabla \cdot \mathsf{P_r} = \boldsymbol{S_r}(\boldsymbol{P}),$$

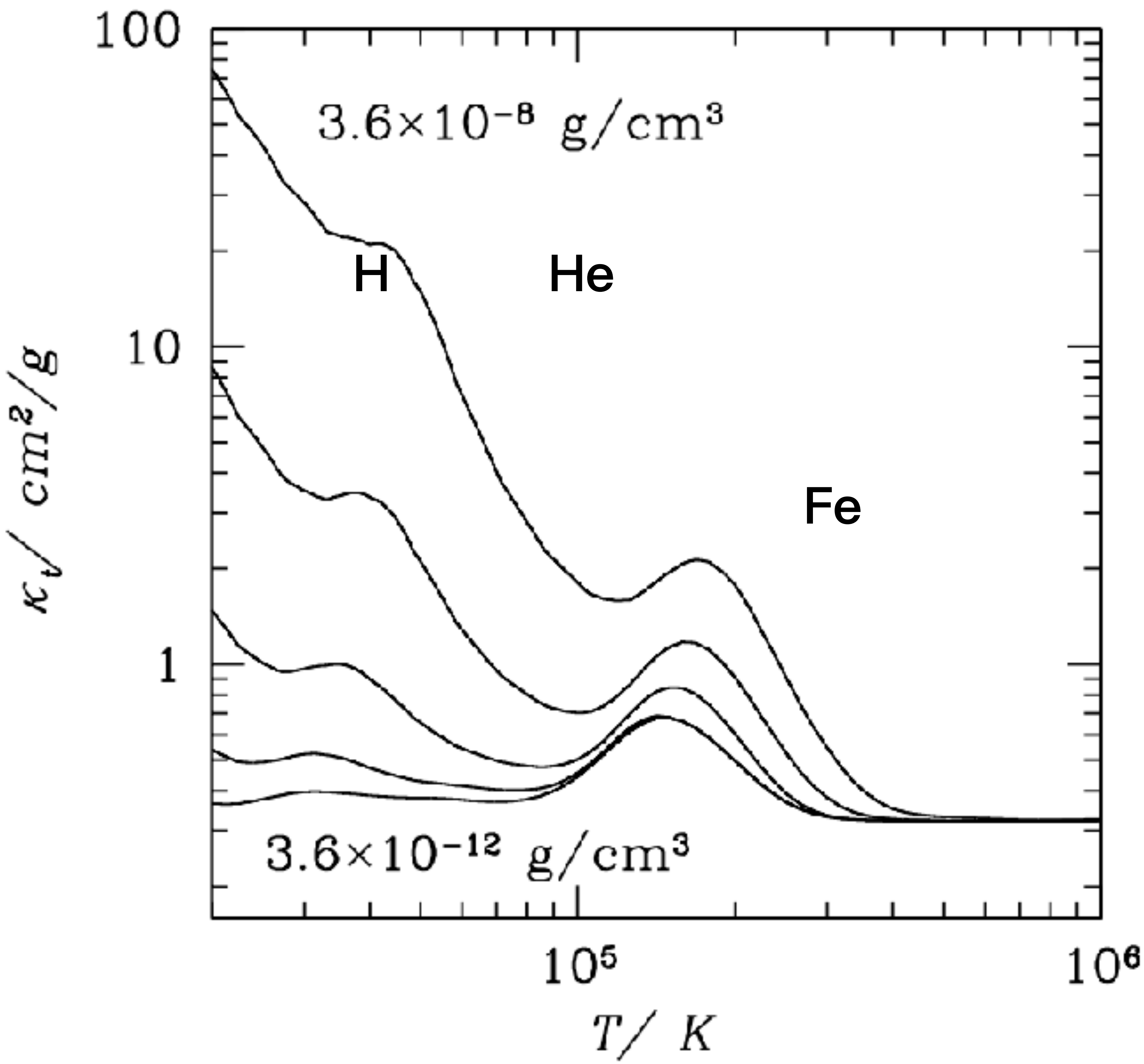
$$F_r = F_{r,0} + v(E_r + P_r)$$

Closure:

$$\frac{\partial I_r}{\partial s} = \kappa_t (S - I_r) \qquad \mathbf{f} = \frac{\int I_r \boldsymbol{nn} d\Omega}{\int I_r d\Omega}, \qquad \mathsf{P_r} = \mathbf{f} E_r$$

# Super-Eddington due to the opacity peaks

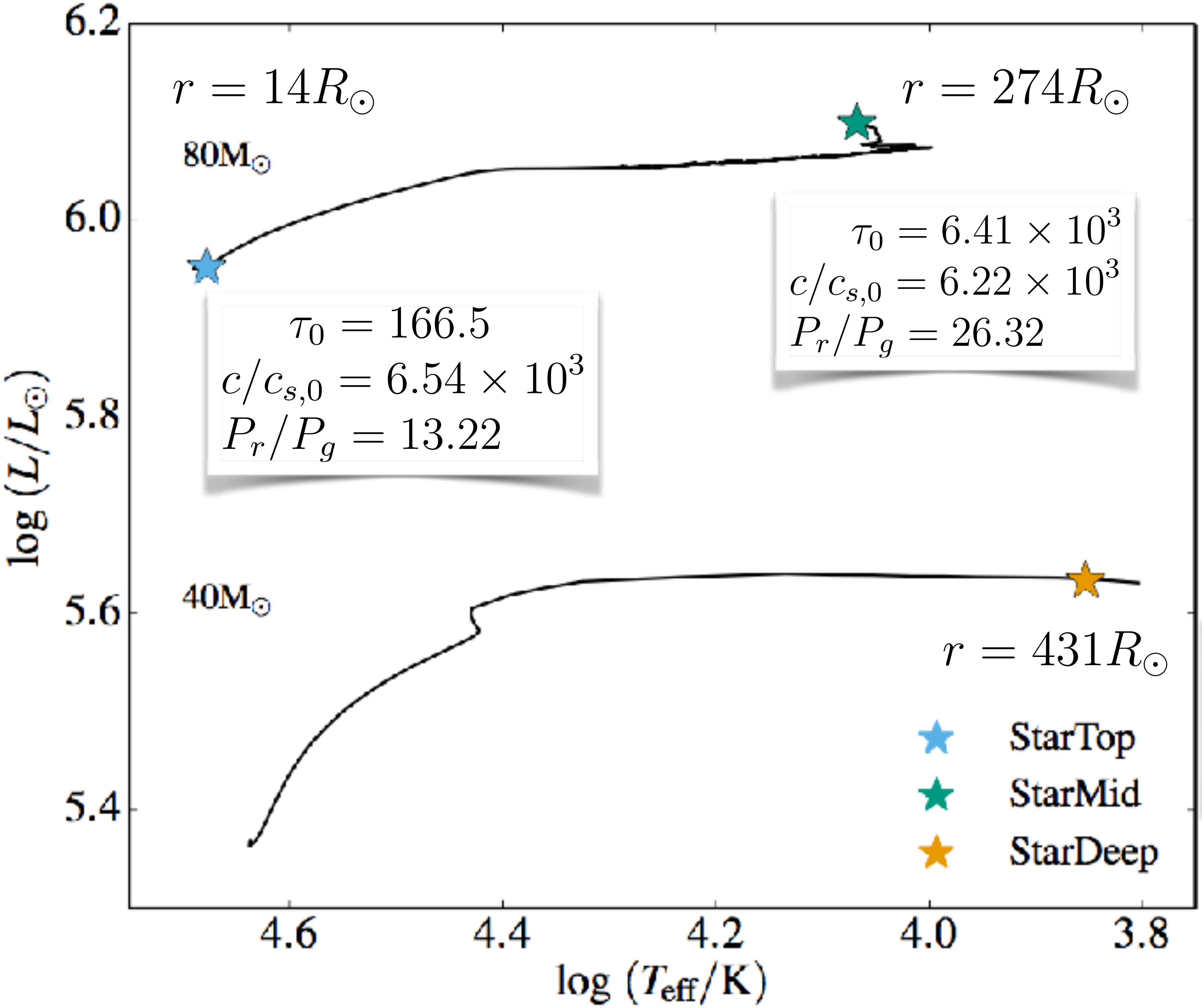
Paxton et al. (2013)
Jiang et al. (2015)




constant density lines

- Advantages of 3D Simulations
  - Capture the radiation (magneto-)hydrodynamic instabilities (convection)
  - Calibrate the 1D mixing length theory in the radiation pressure dominated regime
  - Capture the 3D effects (porosity caused by the density fluctuations)
- Disadvantages of 3D Simulations
  - Cannot cover the whole radial range of the star
  - Cannot evolve for a long time (compared with the life time of the stars)

# The fiducial Models

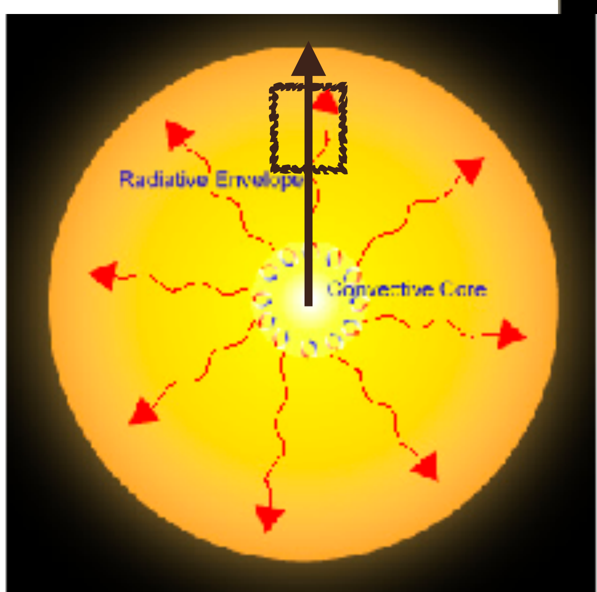



$$\tau_0 = 9.12 \times 10^4$$
$$c/c_{s,0} = 5.99 \times 10^3$$
$$P_r/P_g = 3.96$$

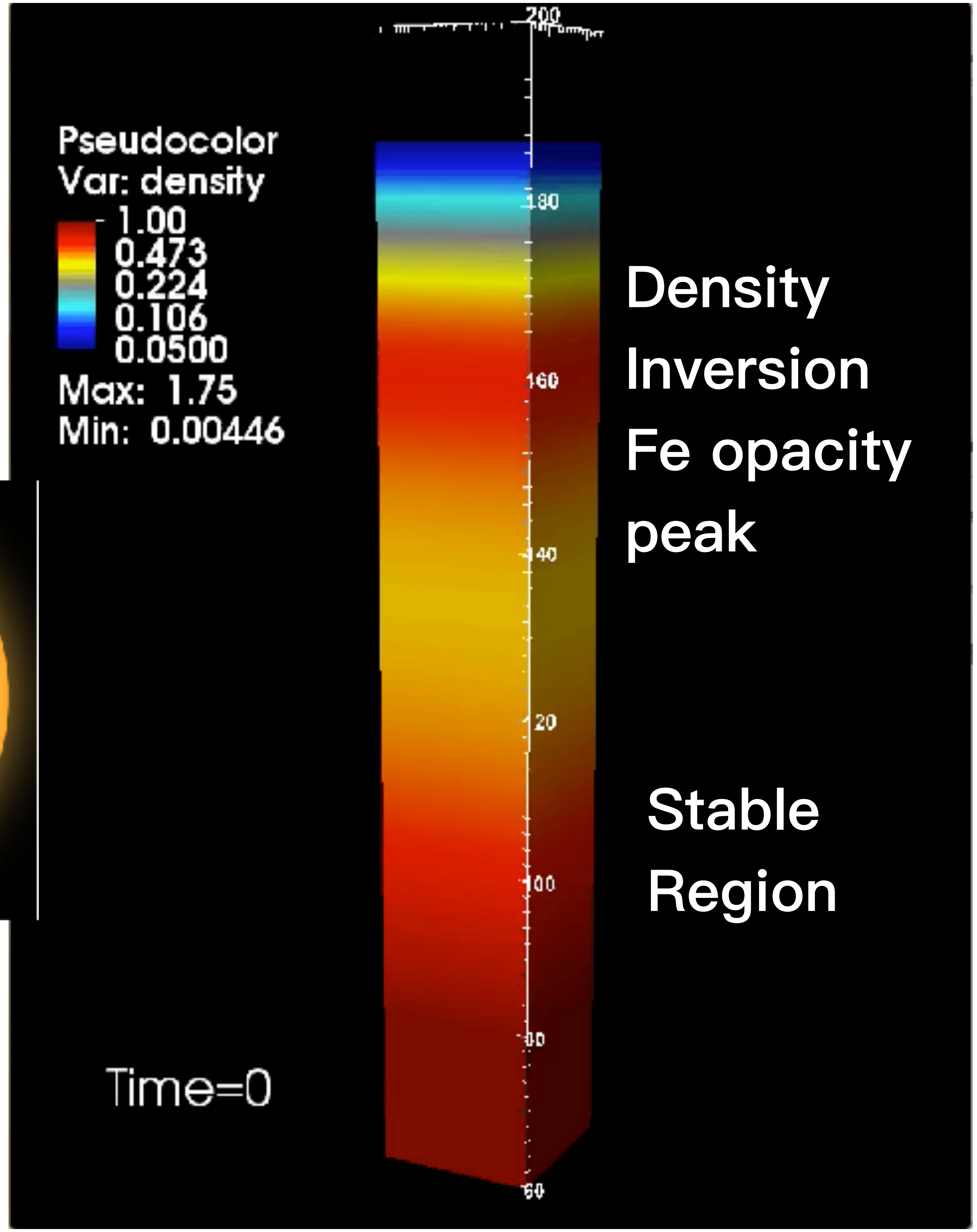
$$c_{s,0} = \sqrt{\frac{a_r T_0^4}{3\rho_0}},$$
$$H_0 = \frac{c_{s,0}^2}{g} = \frac{a_r T_0^4}{3\rho_0 g},$$
$$\tau_0 = \kappa_t(\rho_0, T_0)\rho_0 H_0,$$

# Setup for the local calculations

Open top boundary (Photosphere can be included)




Reflection bottom boundary




Jiang et al. (2015)

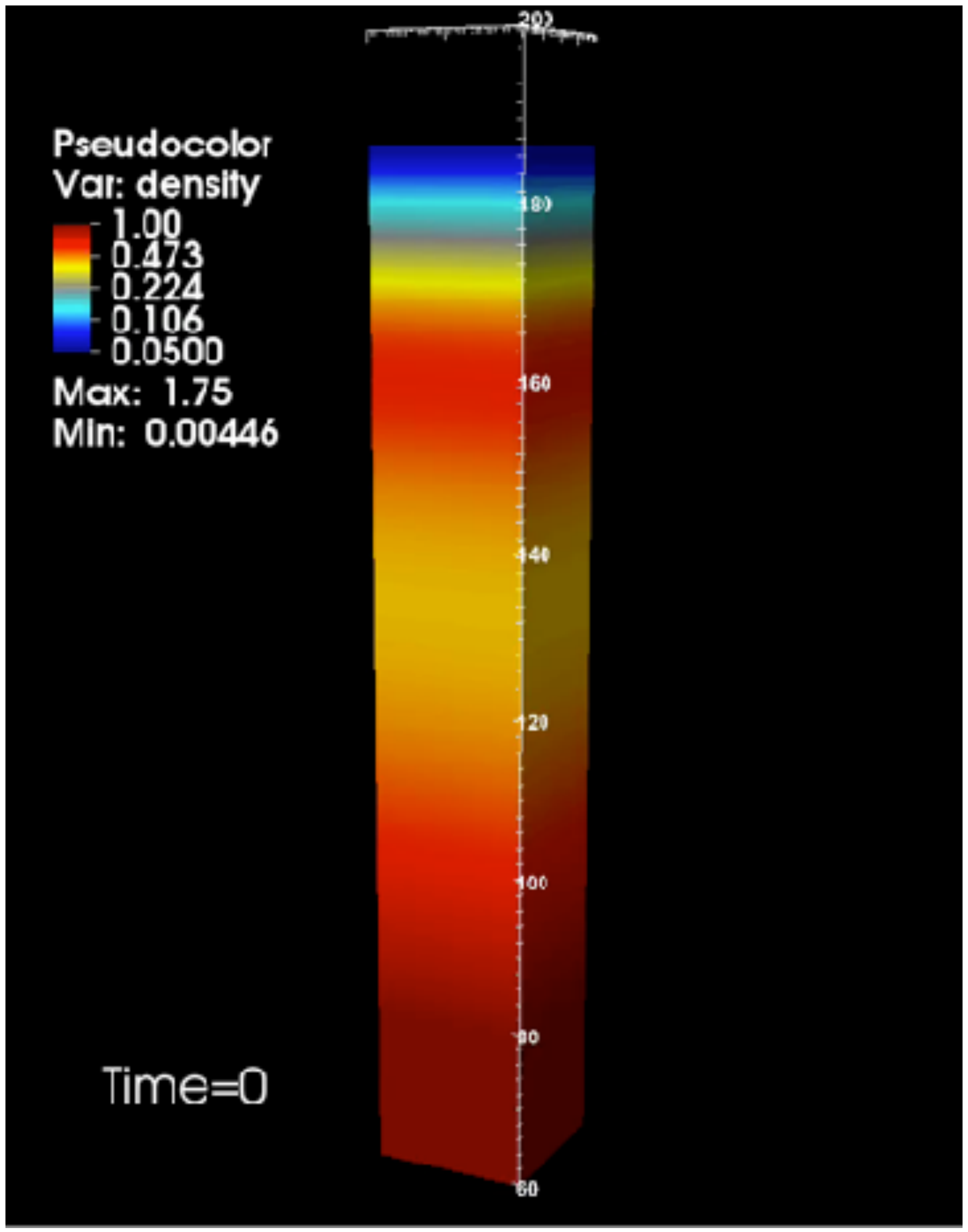
Constant gravity

g

Constant radiation flux coming from the bottom

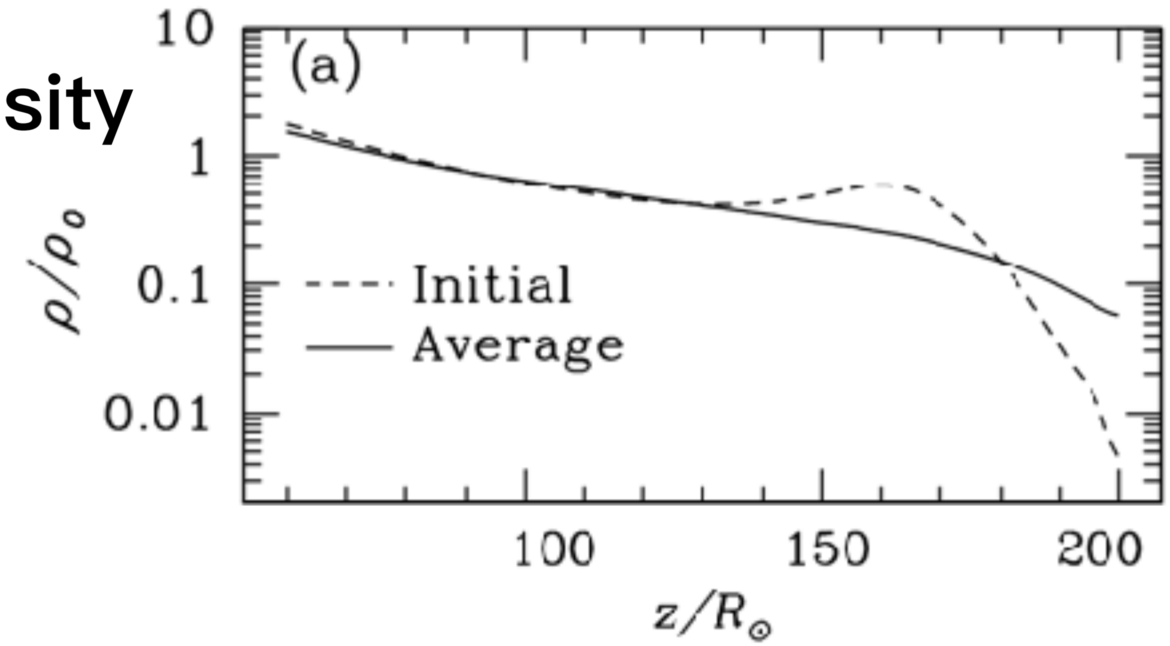
Fr

# 40 Solar Mass YSG: The Case with Efficient Convection




Density

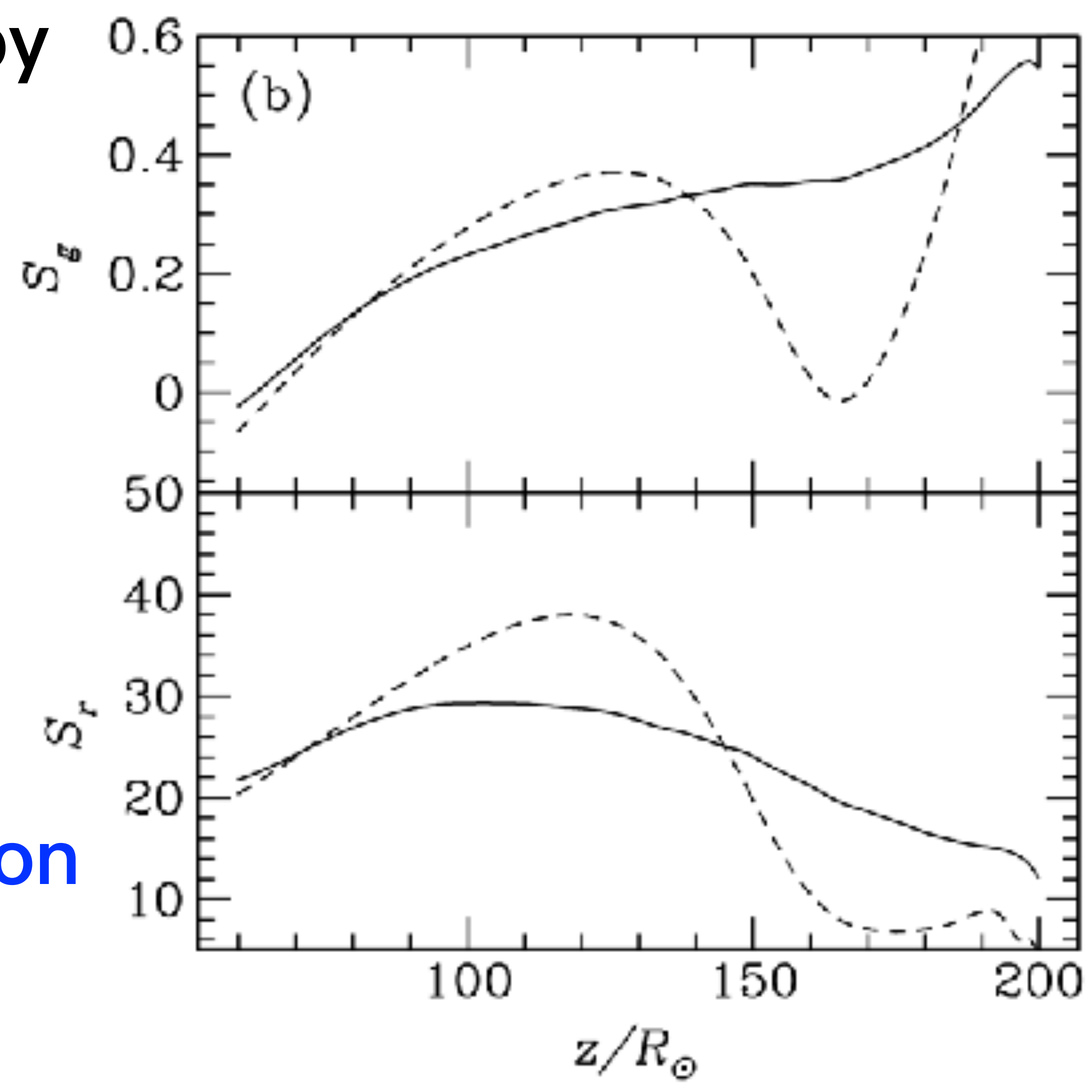
Density



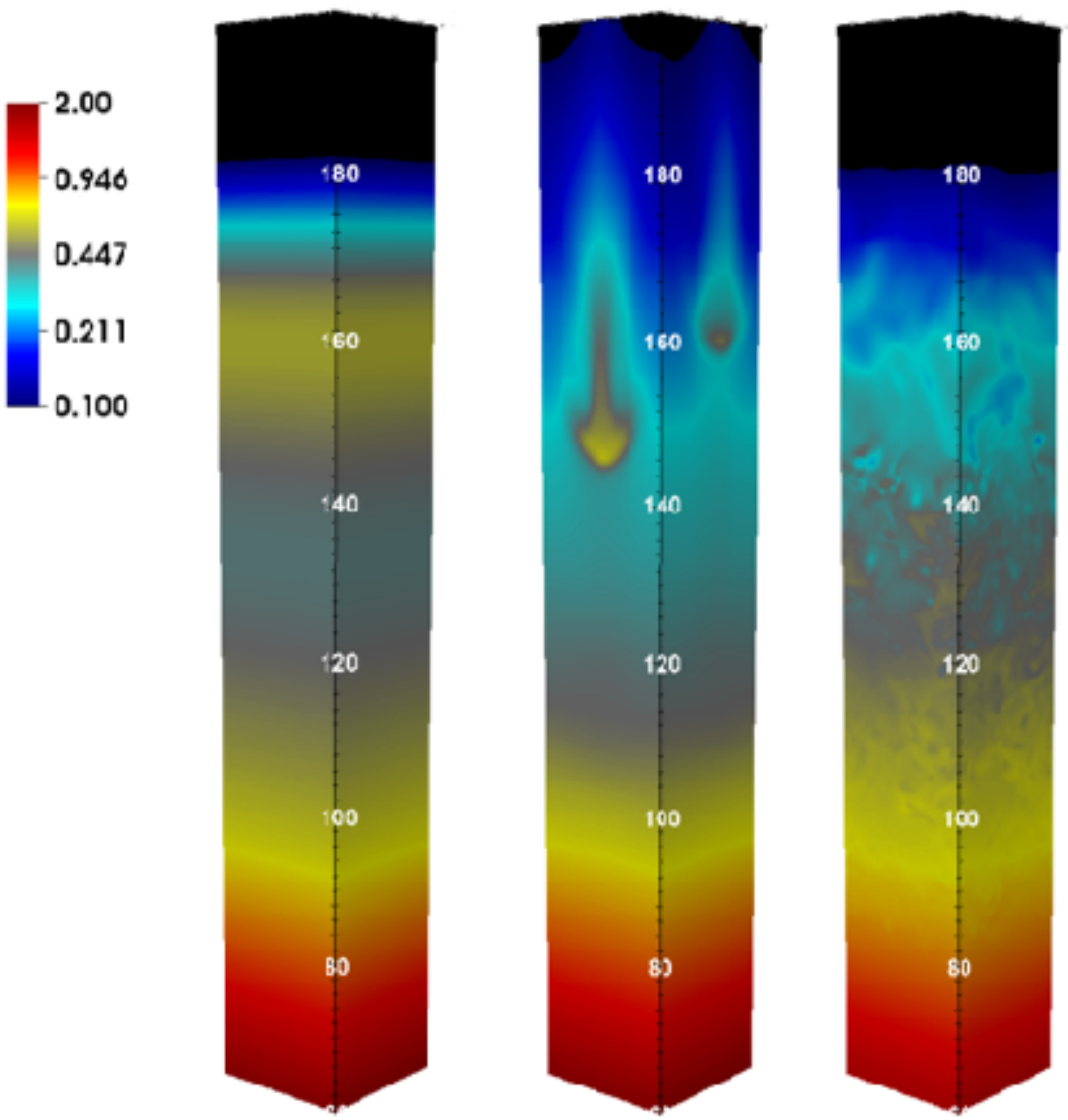

Entropy

Gas

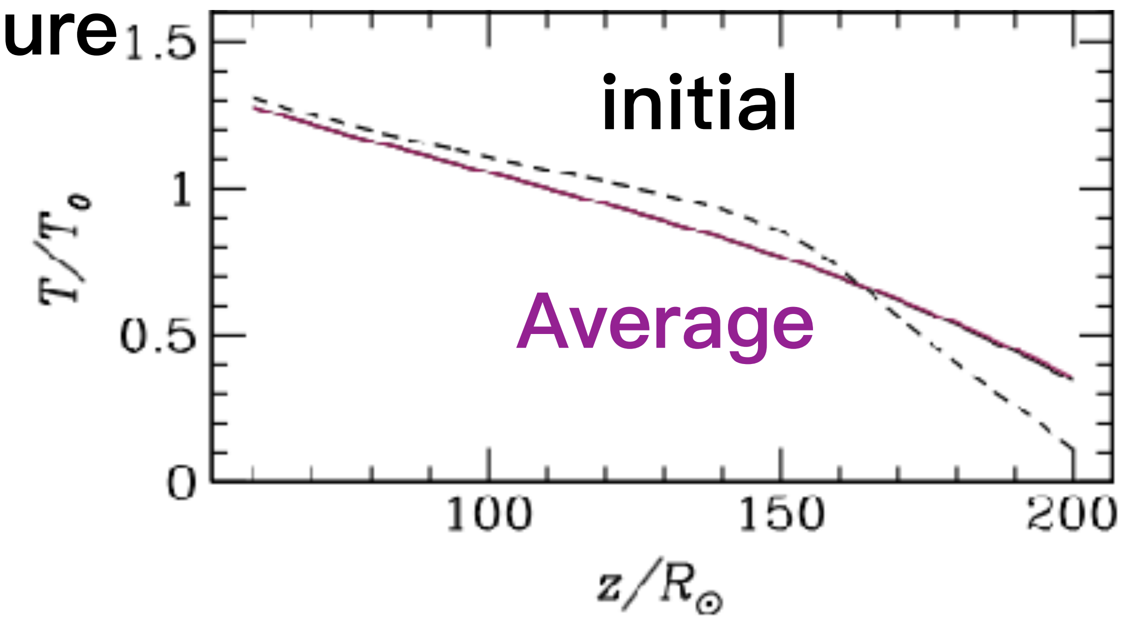
Radiation



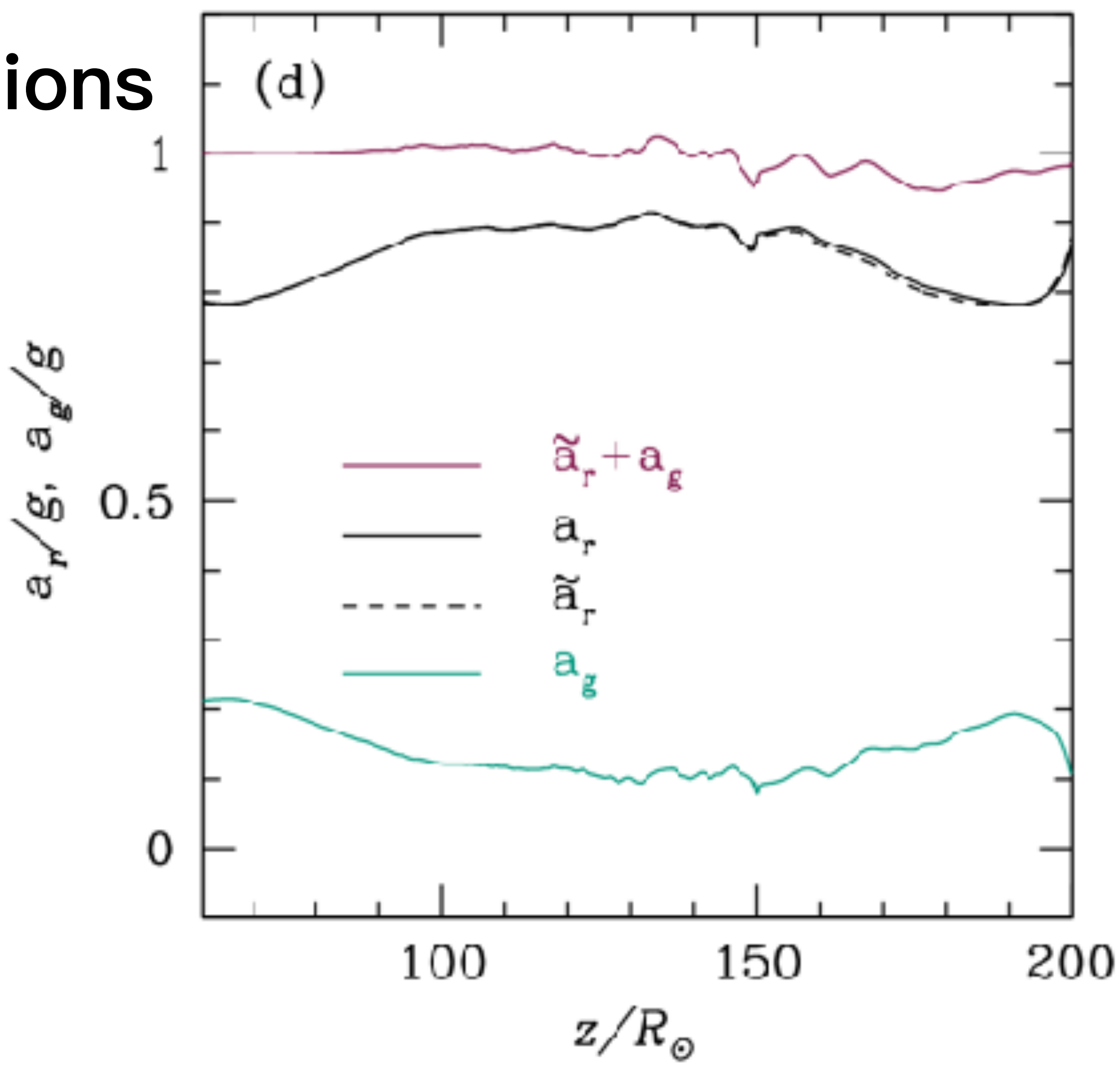
# 40 Solar Mass YSG:
# The Case with Efficient Convection




Temperature



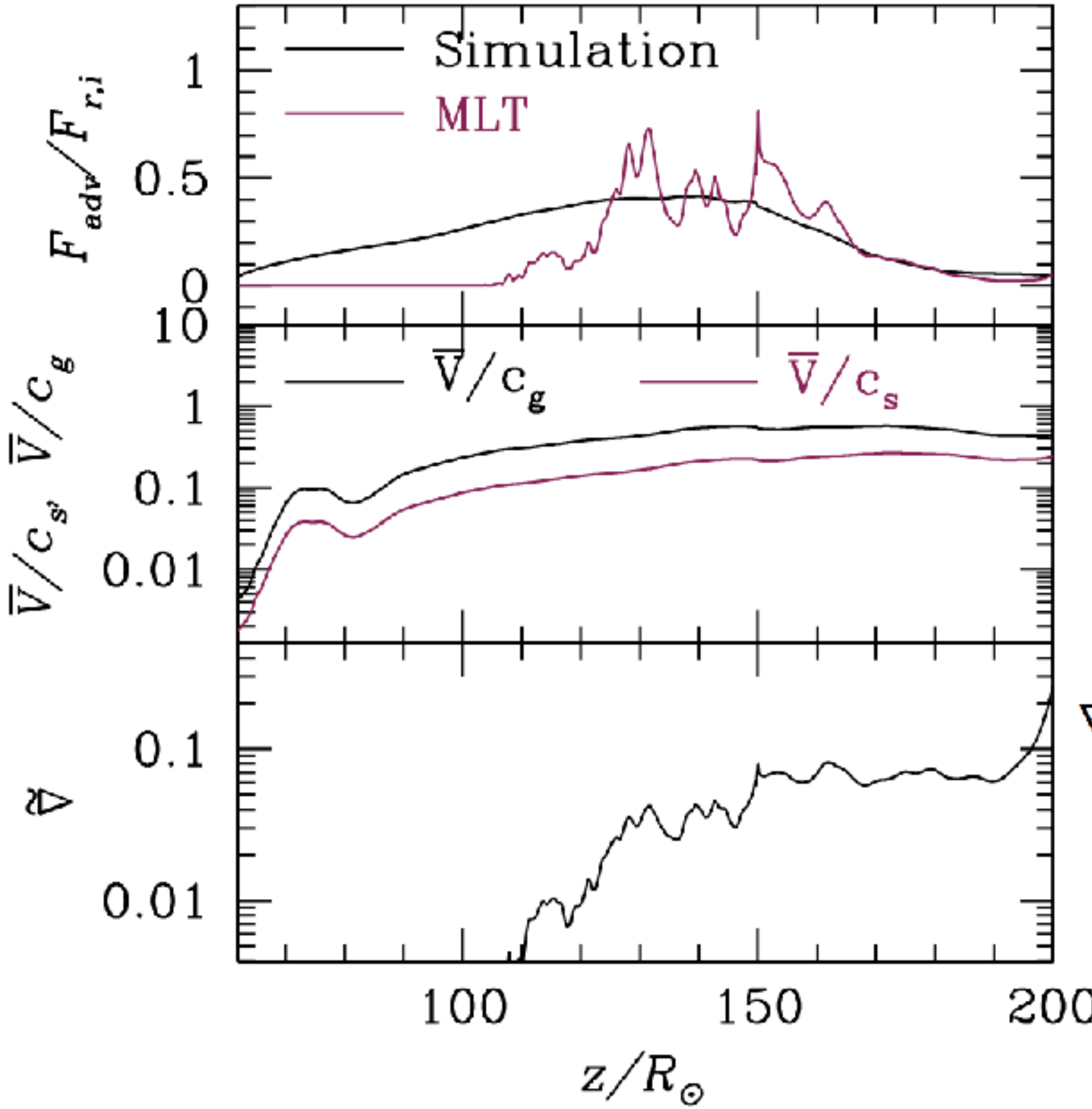

Accelerations




$$\tilde{a}_{\mathrm{r}} = \frac{\langle \rho \kappa_t F_{\mathrm{r,0z}} \rangle}{c \langle \rho \rangle}$$

$$a_{\mathrm{r}} = \frac{\kappa_t}{c} F_{\mathrm{r,0z}}.$$

# Compared with MLT




$\alpha = 0.55$

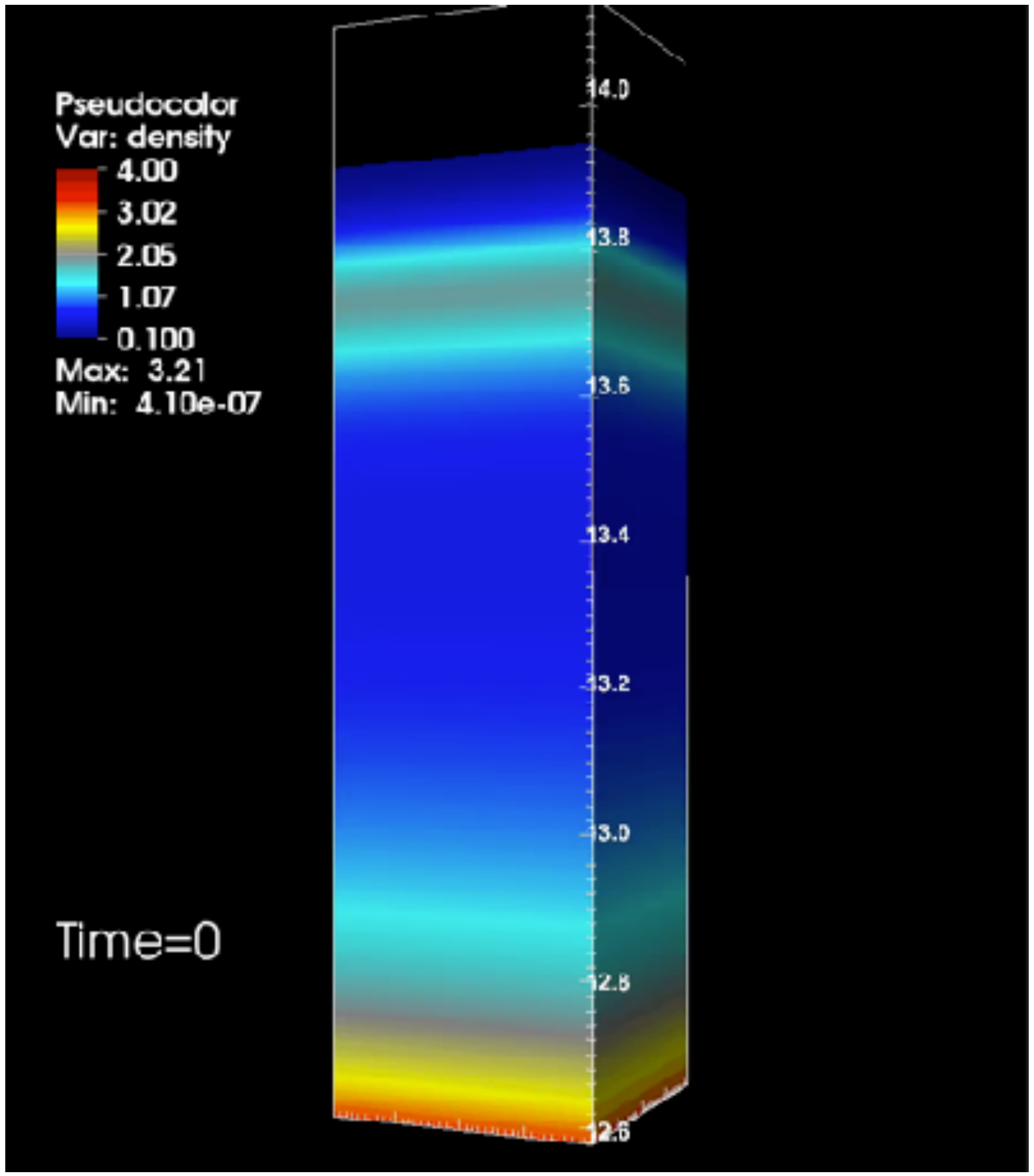
Subsonic convection.

$\nabla \equiv d \ln T / d \ln(P + P_r)$

$\tilde{\nabla} \equiv (\nabla - \nabla_{ad})/\nabla_{ad}$

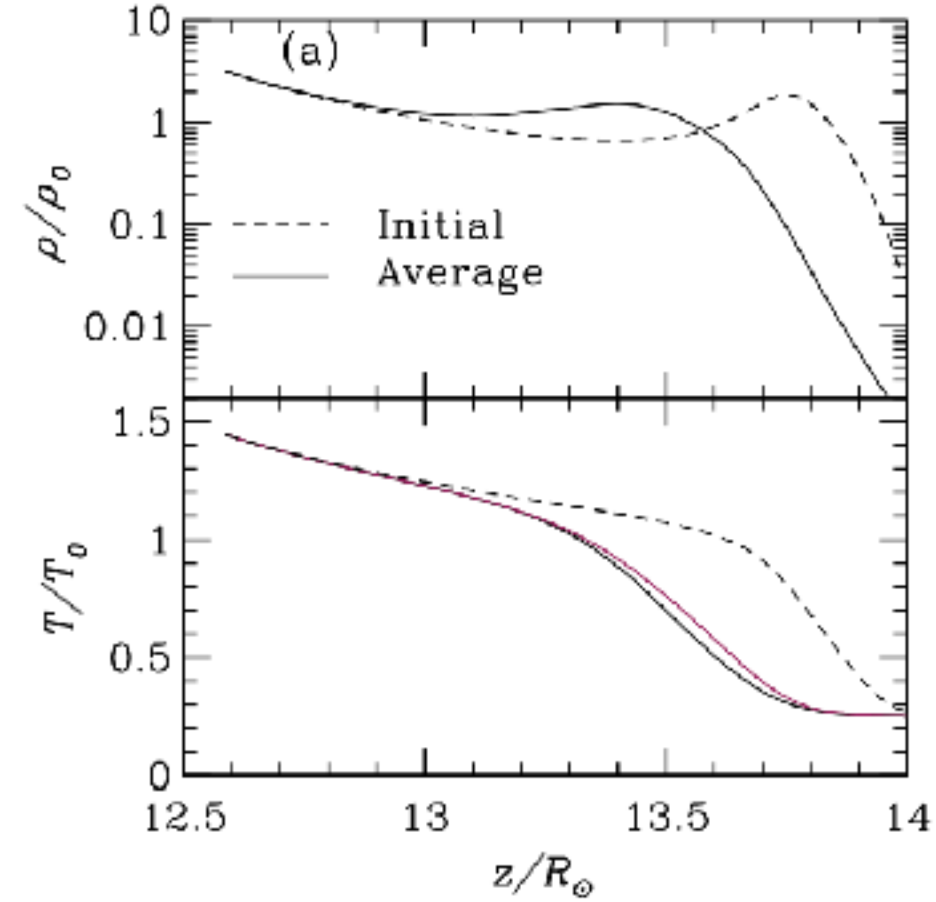
Close to adiabatic.

# 80 Solar Mass ZAMS:
# The Case with Inefficient Convection



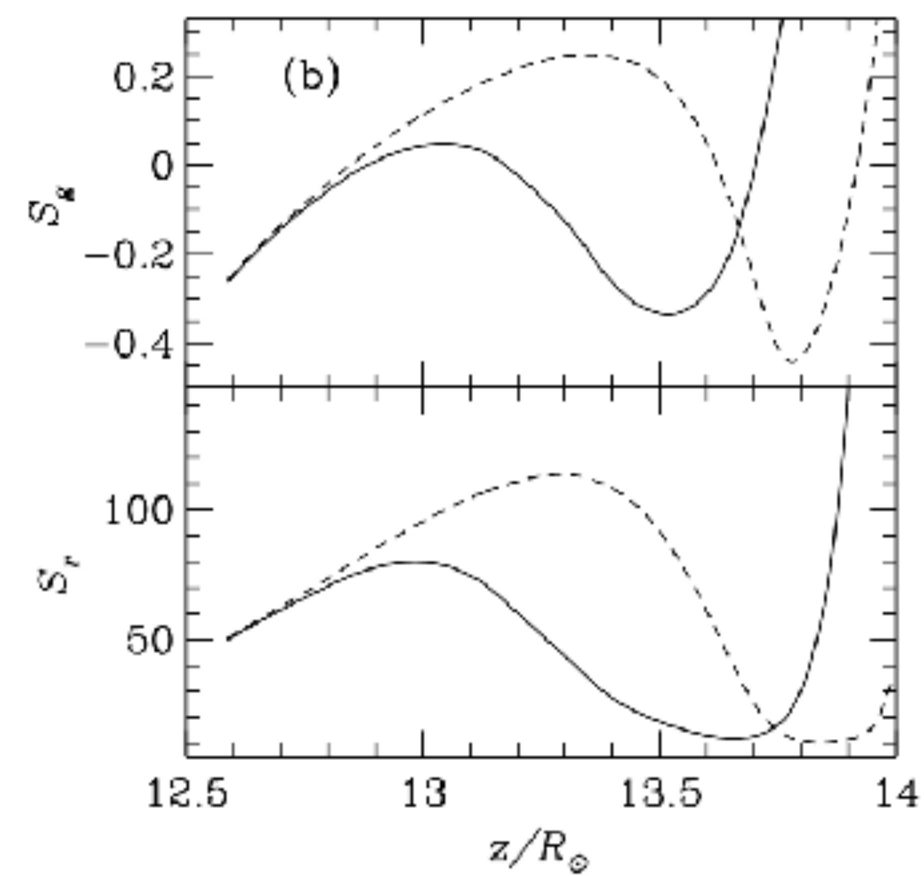

Density

Density

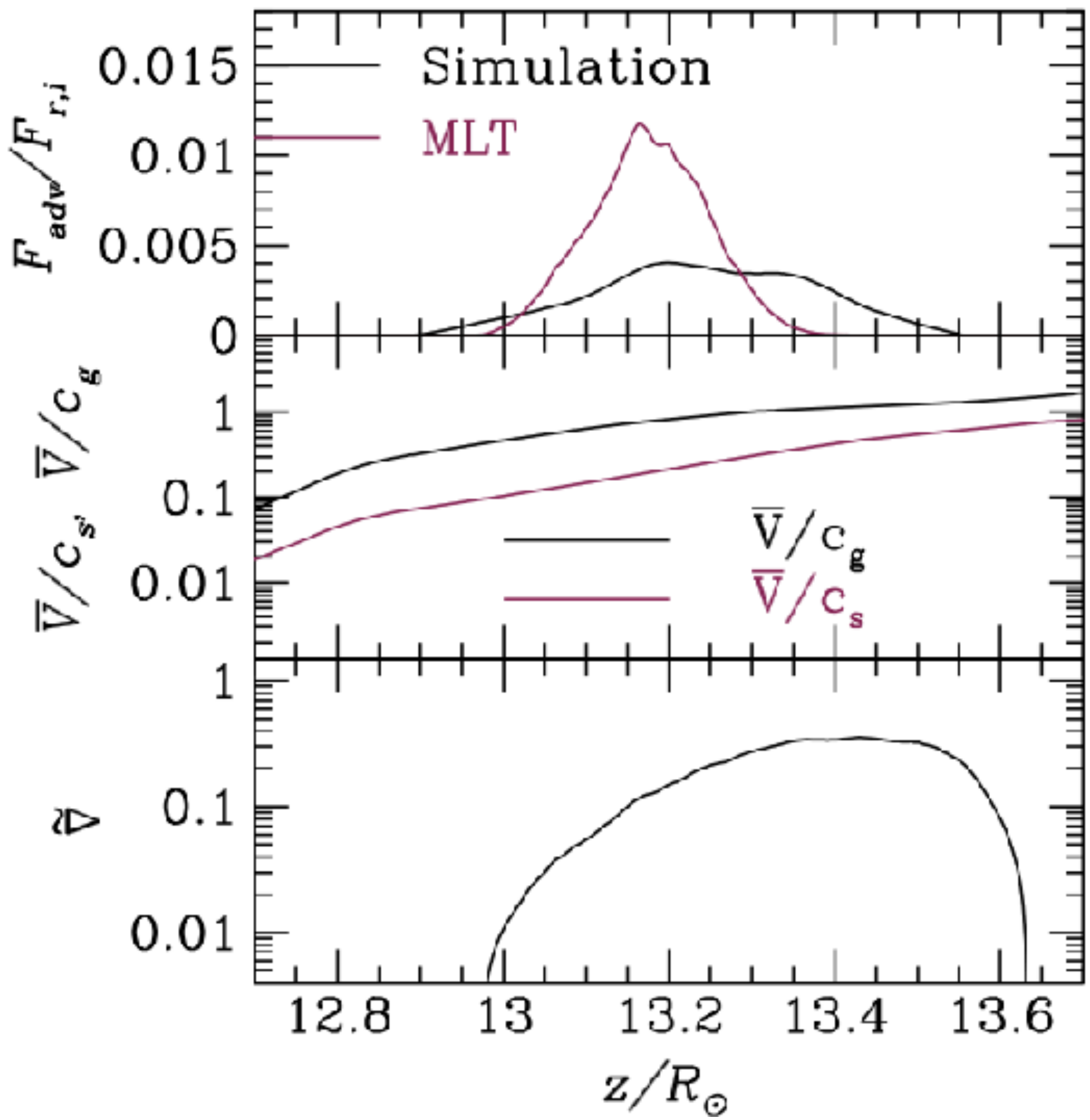



Temperature

Entropy



# 80 Solar Mass ZAMS:
# The Case with Inefficient Convection

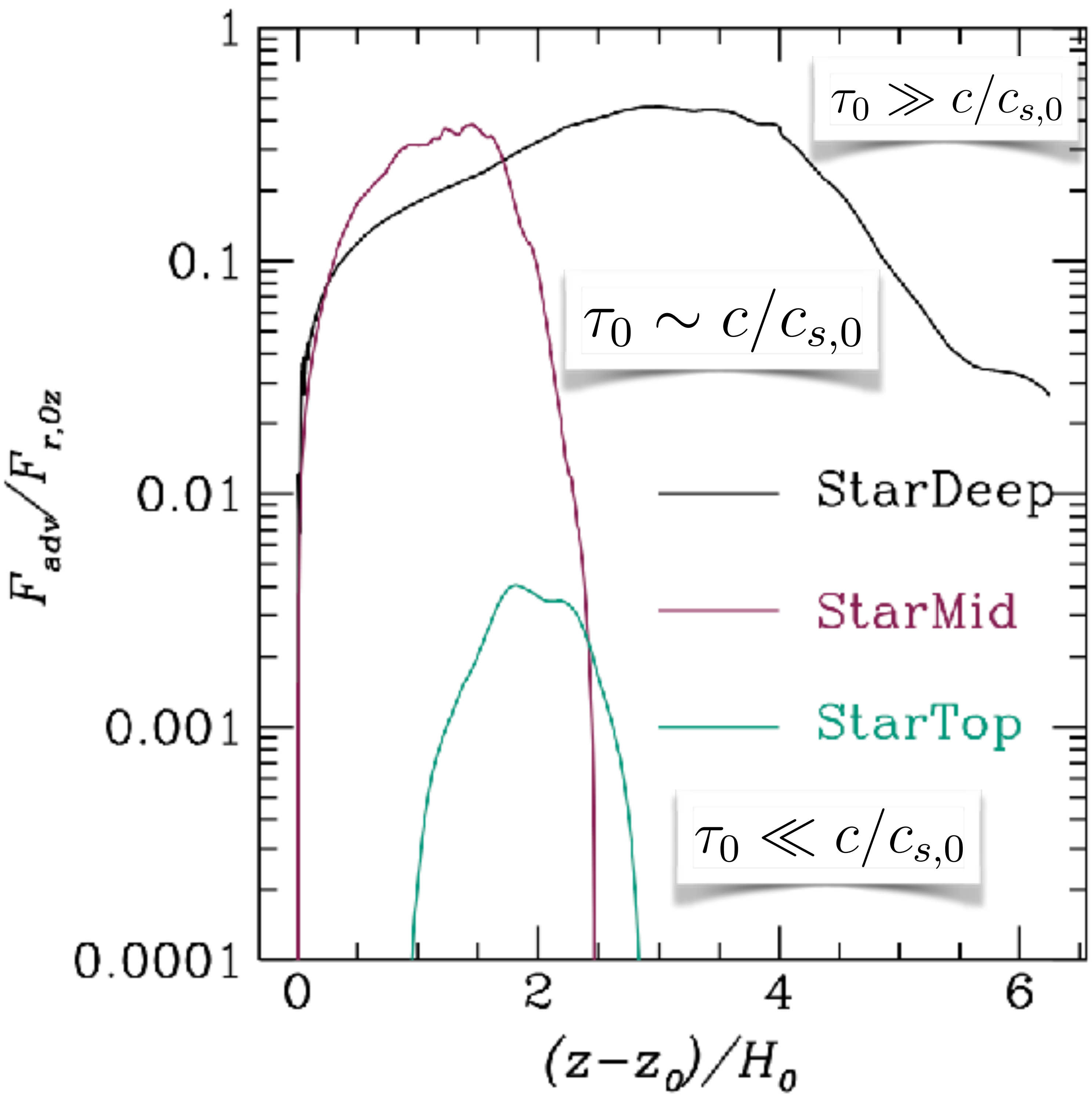



Convection flux much smaller than the MLT predicted value

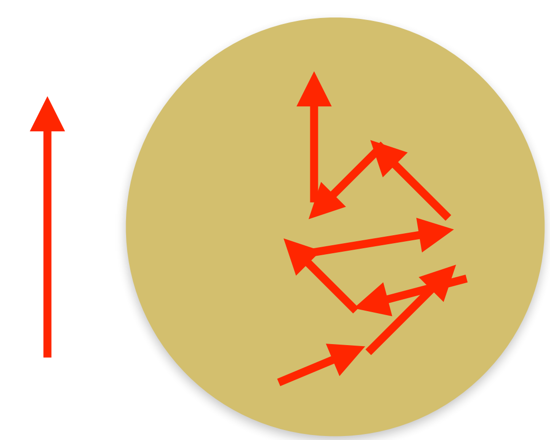
Supersonic turbulent velocity > 50 km/s

Larger difference compared to the adiabatic value.

# Summary of Convection in Radiation Pressure dominated regime




**The competition between diffusion time scale and advection time scale.**

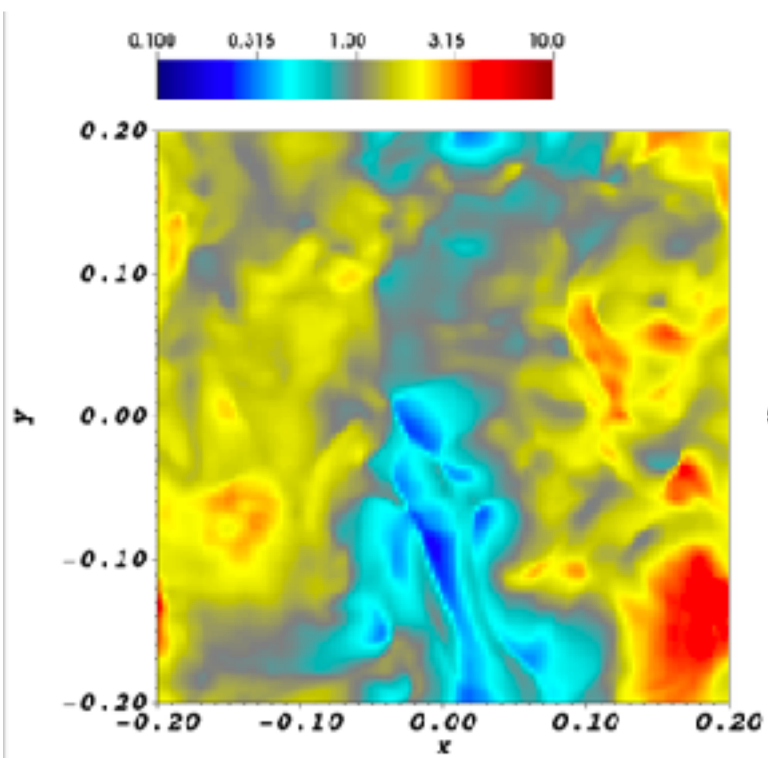


# The Porosity Factor When $\tau_0 \ll c/c_{s,0}$

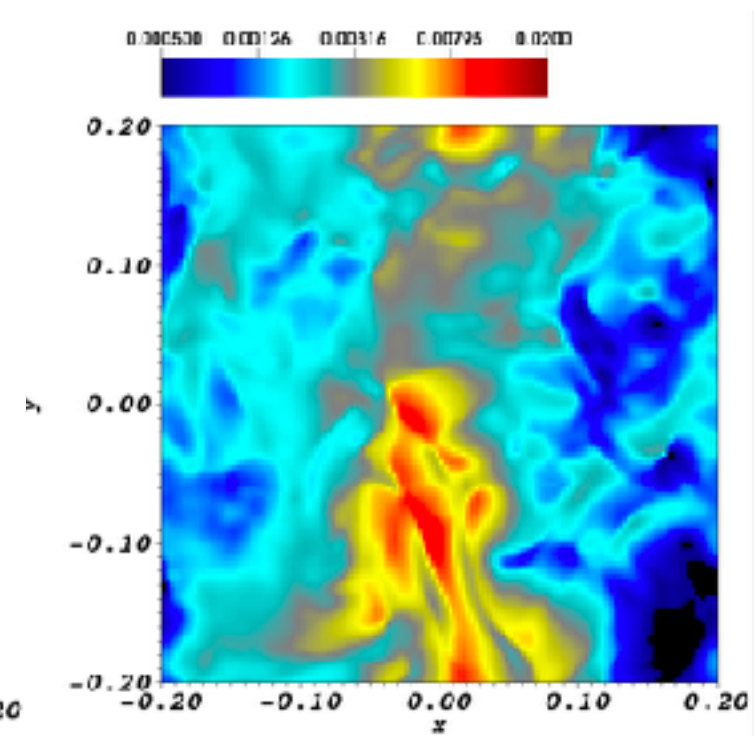
Shaviv (1998)

Van Marle et al. (2008)

Horizontal slice

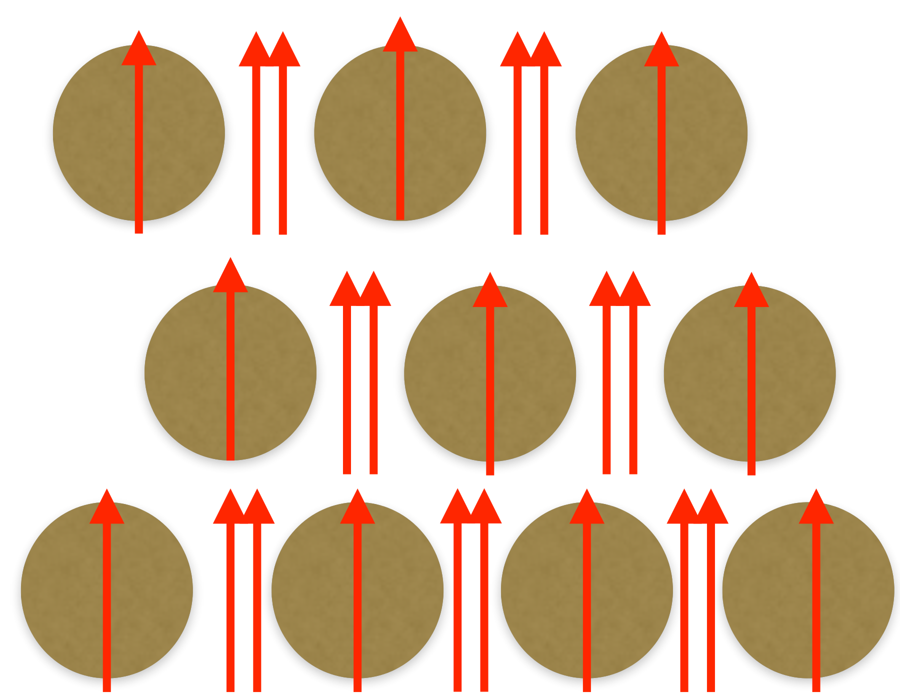



Density




Radiation Flux

$$\tilde{a}_r = \frac{\langle \rho \kappa_t F_{r,0z} \rangle}{c \langle \rho \rangle} \qquad a_r = \frac{\langle \kappa_t F_{r,0z} \rangle}{c}$$



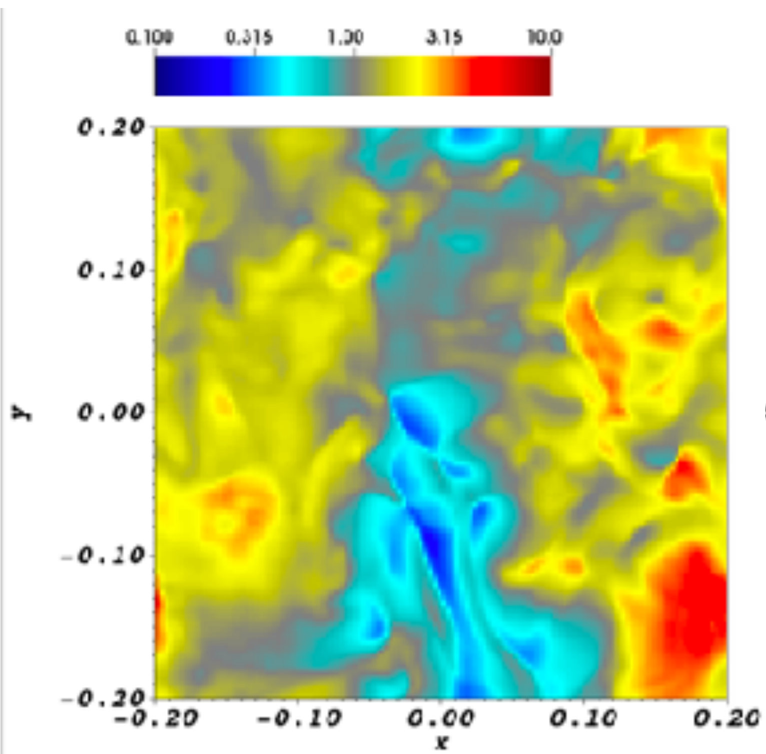
Vertical Structure

# The Porosity Factor When $\tau_0 \ll c/c_{s,0}$

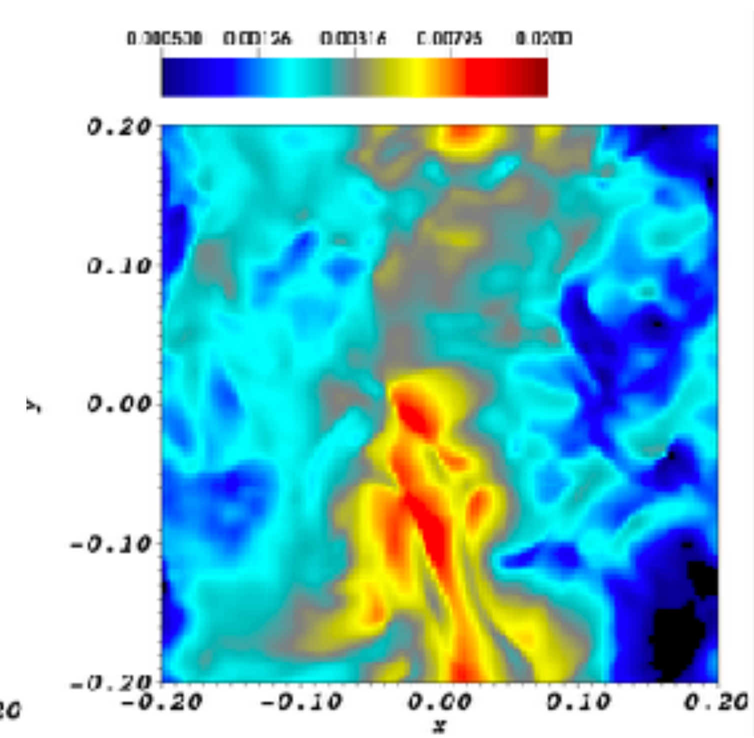
**Horizontal slice**

**Owocki (2014)**

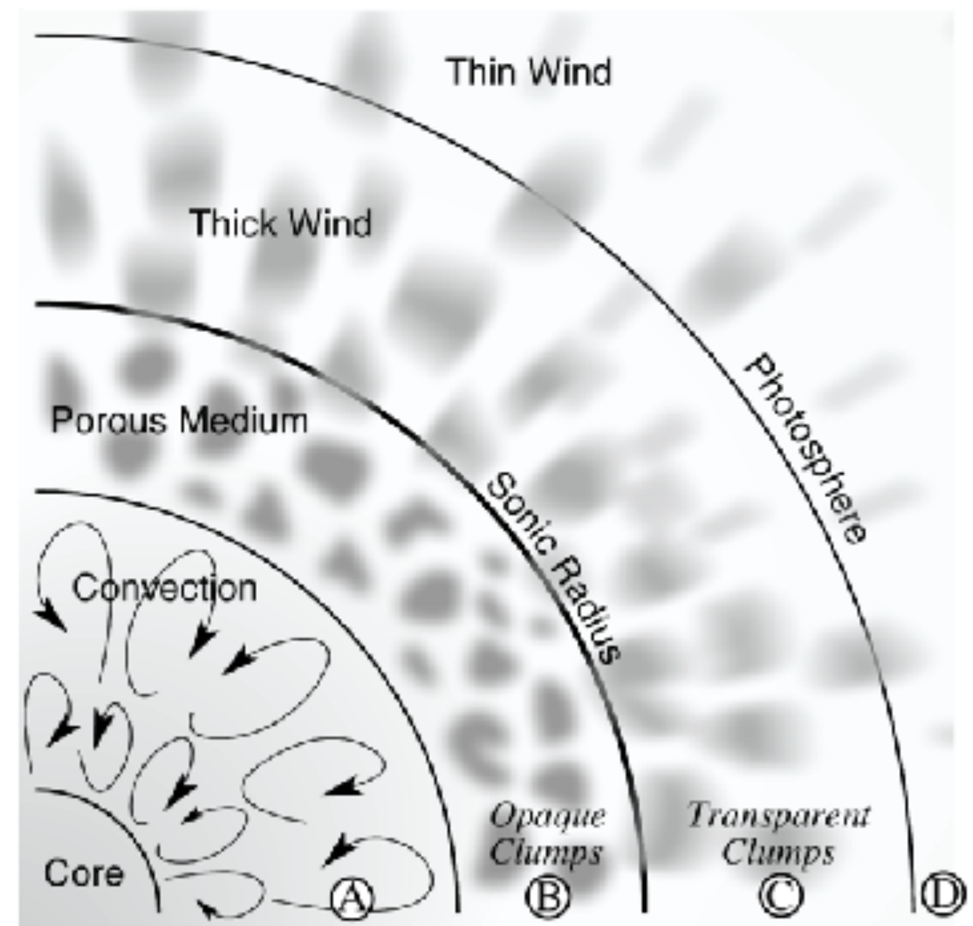
**See Stan's talk**




**Density**

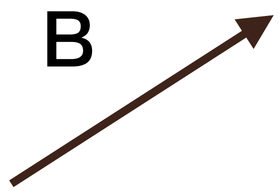



**Radiation Flux**

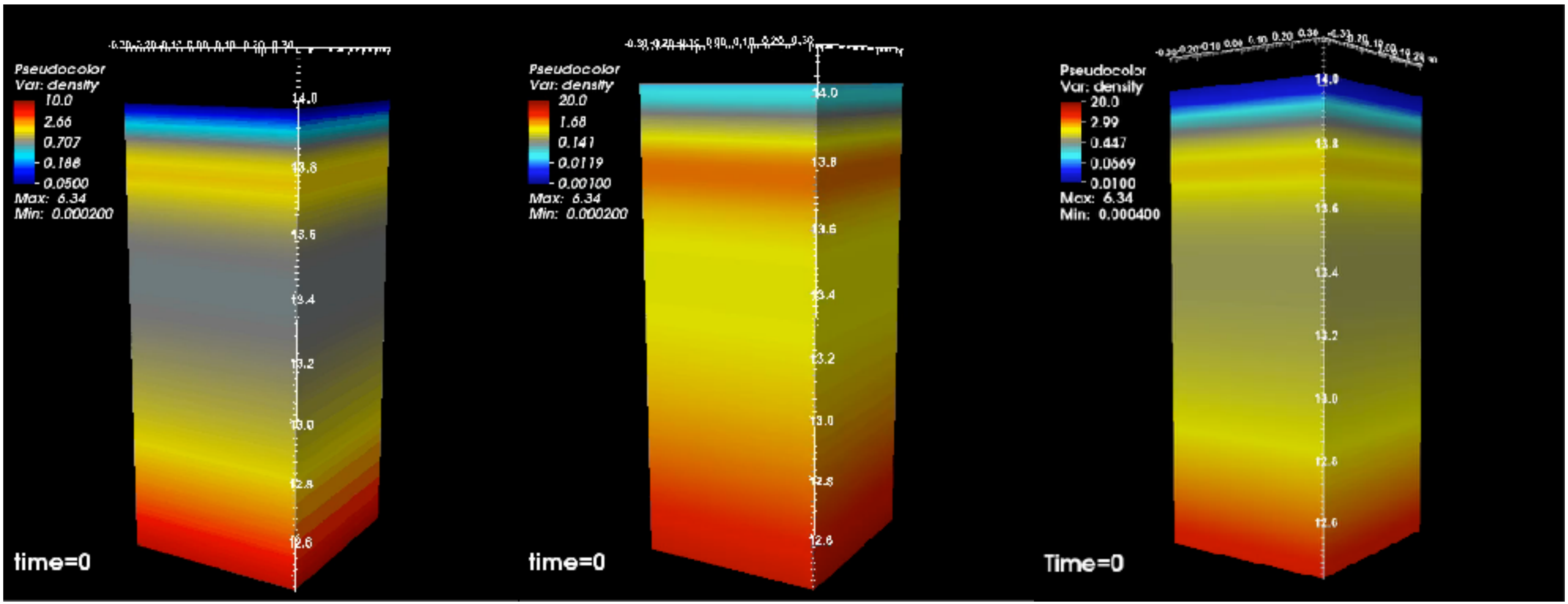



$$\tilde{a}_r = \frac{\langle \rho \kappa_t F_{r,0z} \rangle}{c \langle \rho \rangle} \qquad a_r = \frac{\langle \kappa_t F_{r,0z} \rangle}{c}$$

# Effects of Magnetic Fields




Jiang et al. (2017, arXiv:1612.06434)

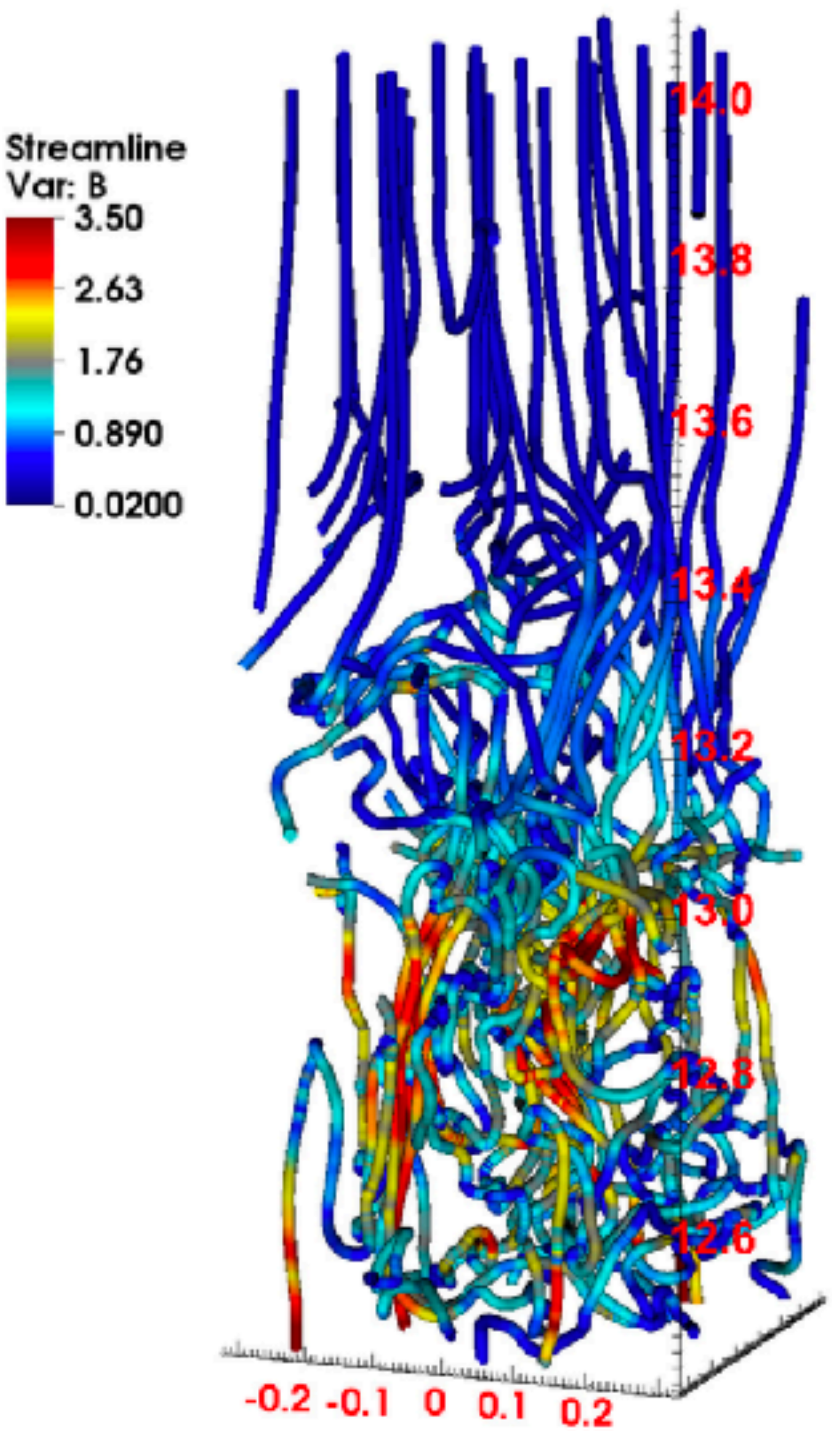



$B_{z,0} = 60G$ $B_{y,0} = 121G$

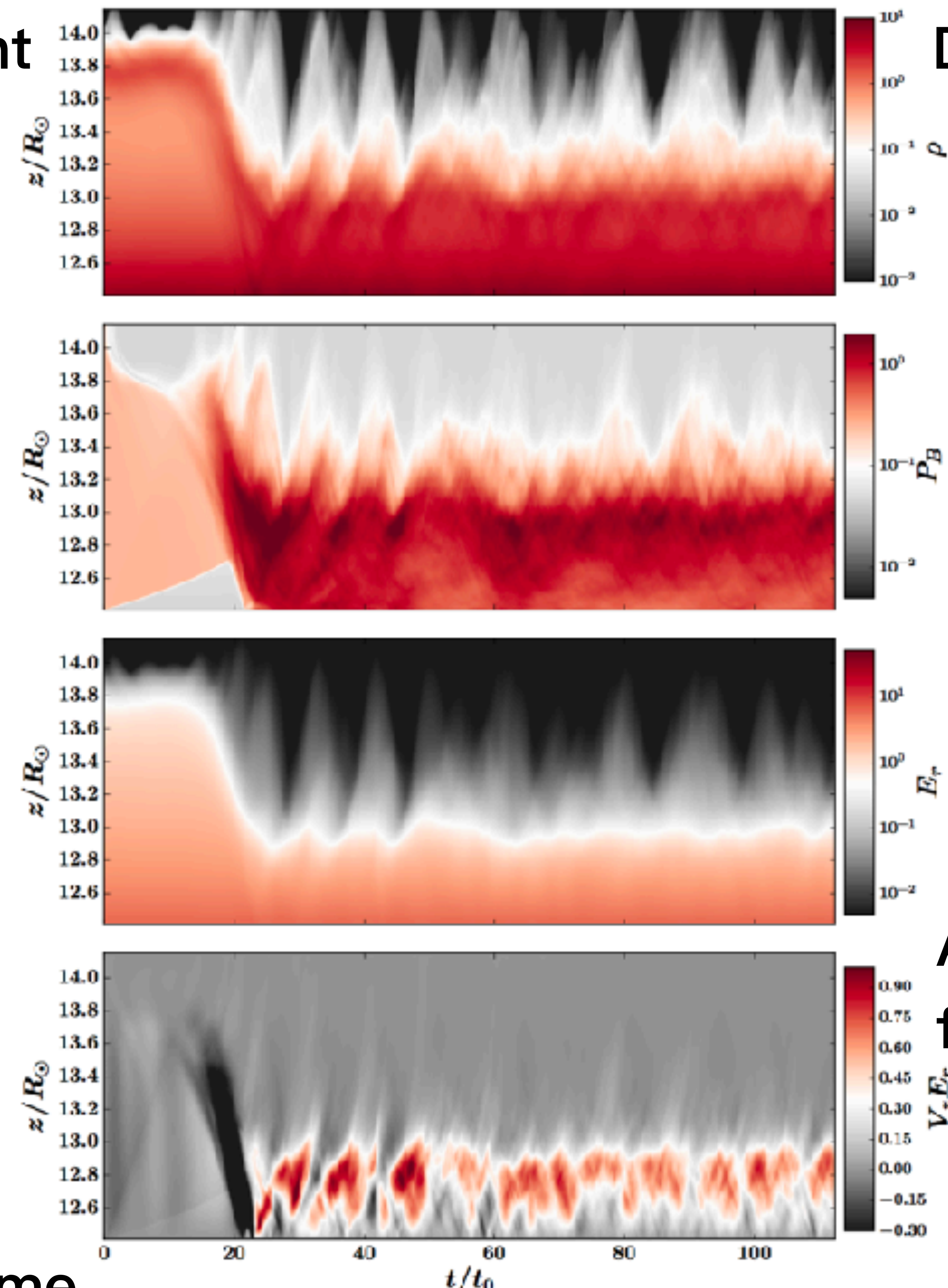
$B_{z,0} = 382G$ $B_{y,0} = 764G$

$B_{z,0} = 382G$ $B_{y,0} = 3819G$

# Magnetic Fields Amplified by the Convection




Height




Density

Magnetic pressure

Radiation Energy

Advection flux

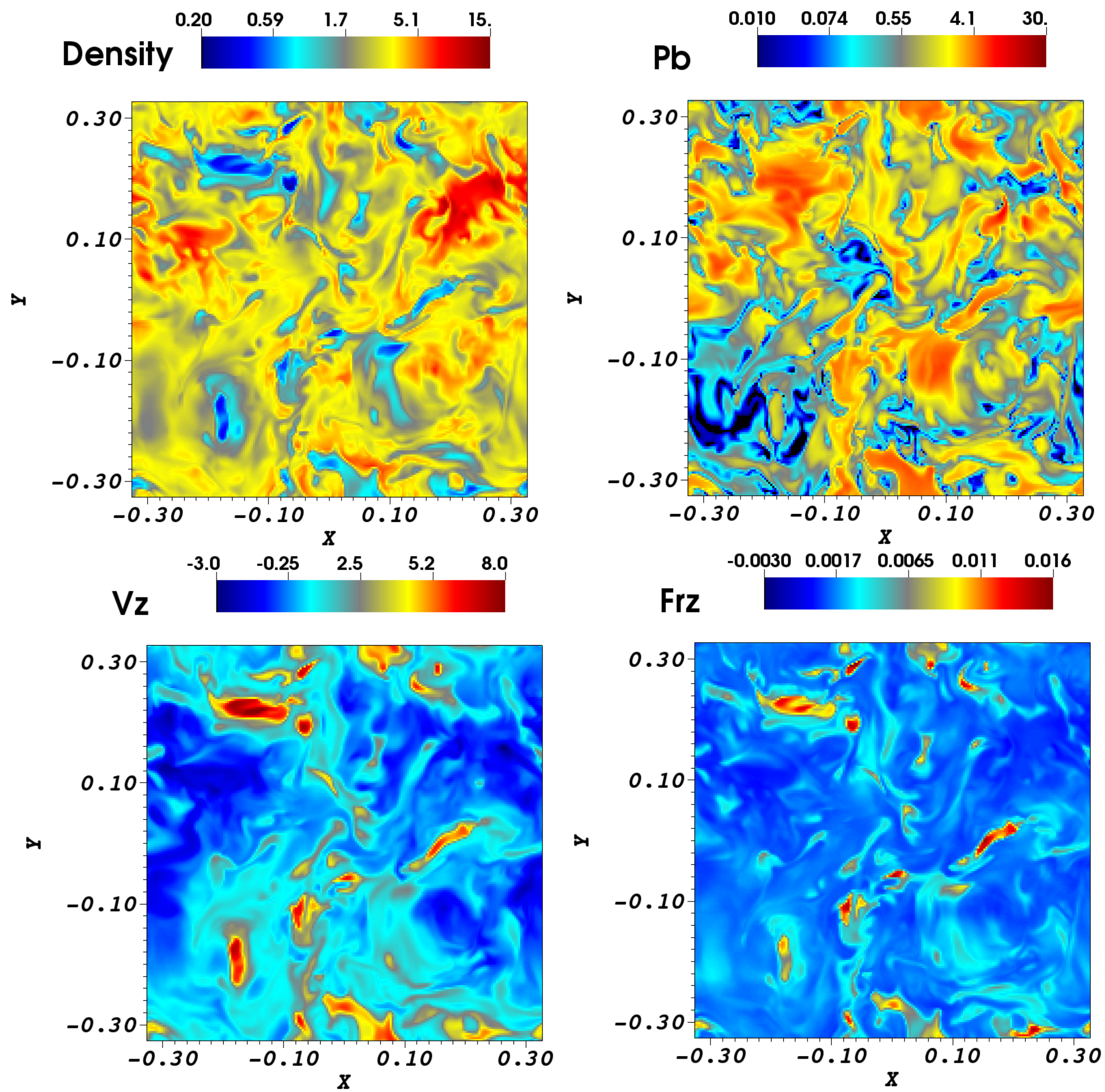
Time

# Magnetic Fields Increase Density Fluctuations

Horizontal slice at

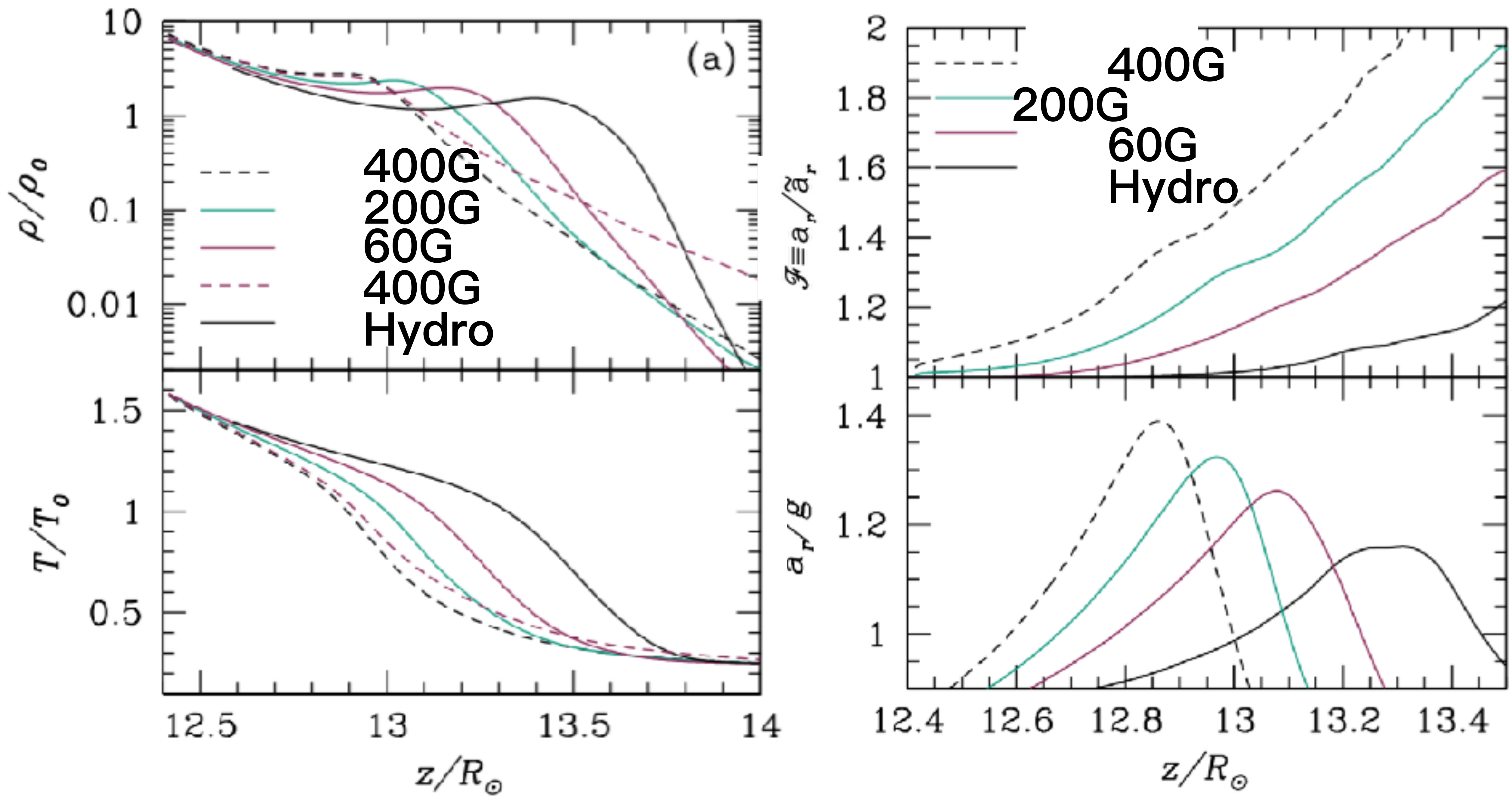
$$z = 12.9 r_{\odot}$$

- Magnetic fields increase the density fluctuations, and the porosity factor.
- Magnetic buoyancy increases the advective flux.

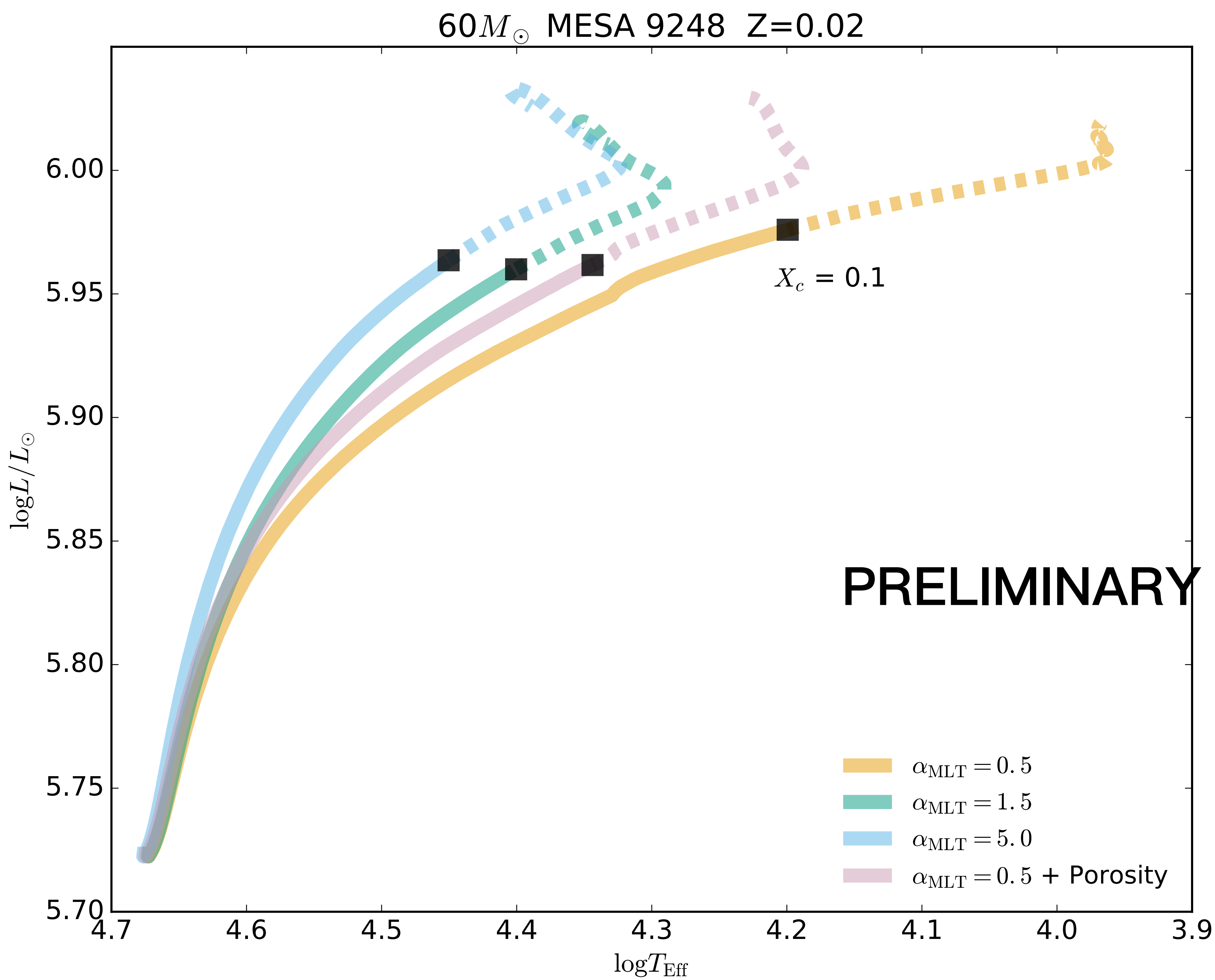


# Effects of Magnetic Field

Porosity factor



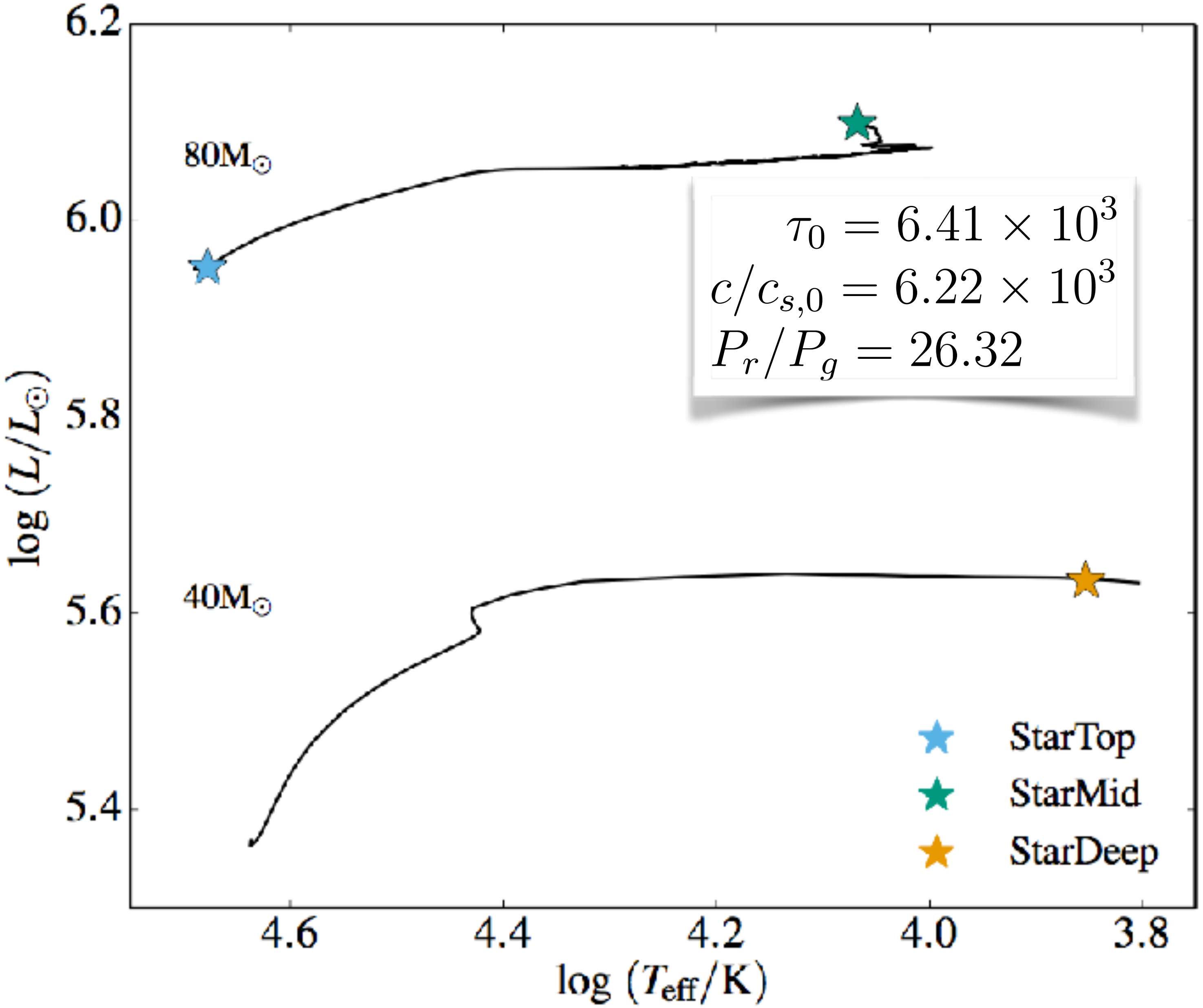
# 1D models with and without Porosity

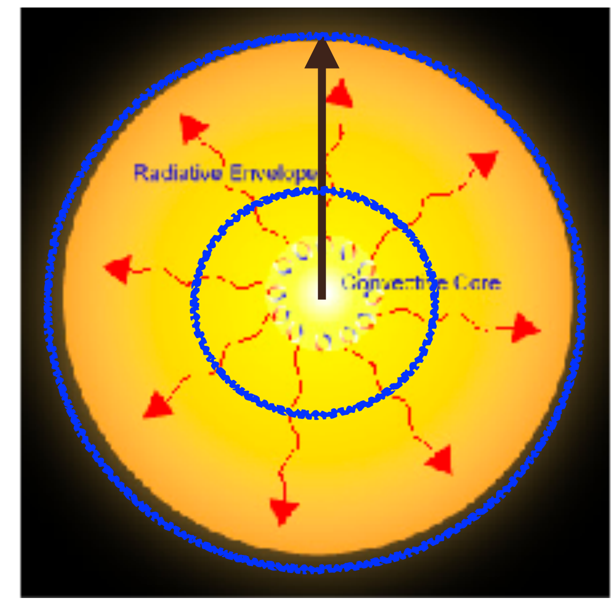



PRELIMINARY RESULTS!

With Matteo Cantiello

# Global Structures of the Massive Star Envelopes

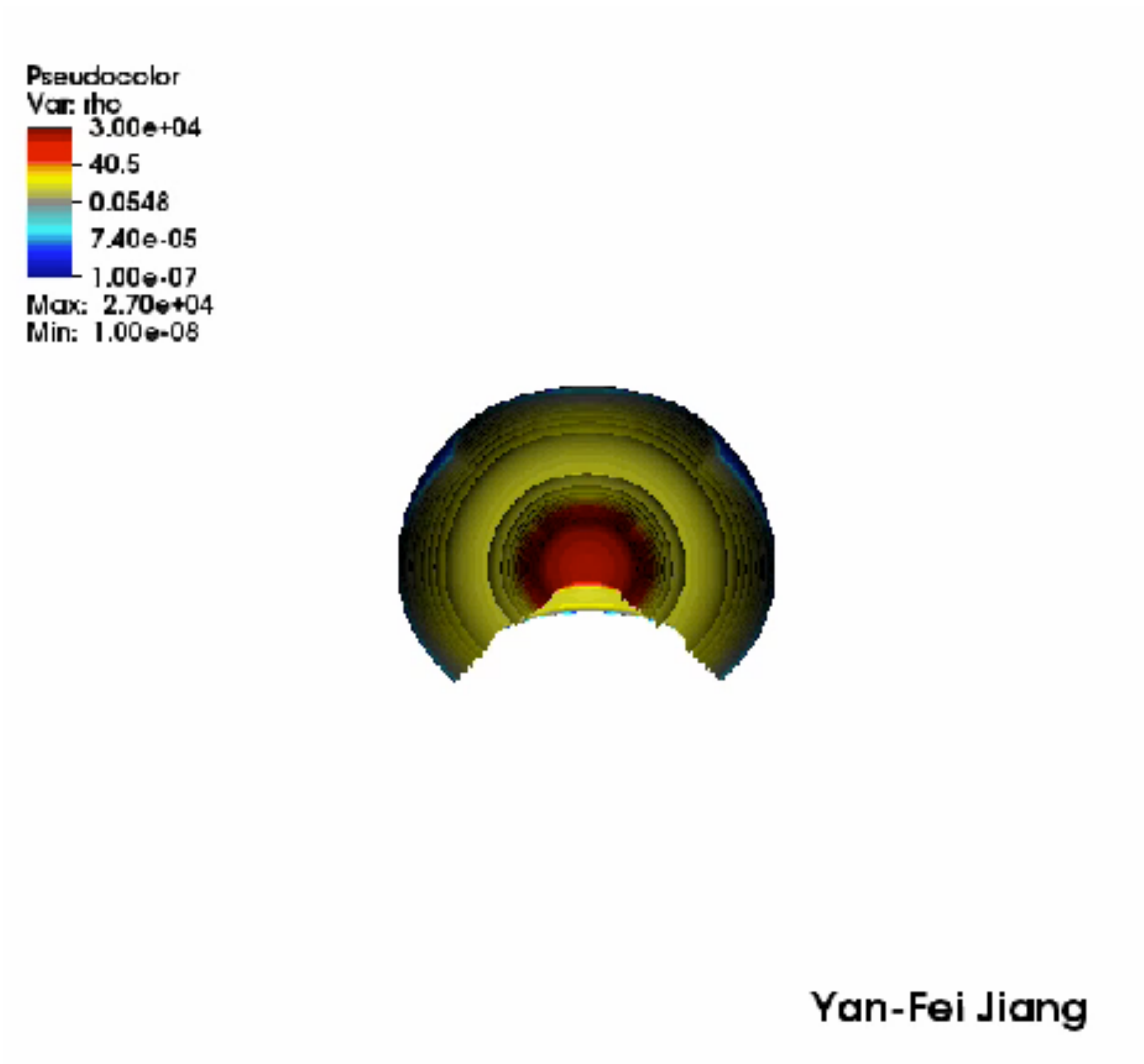


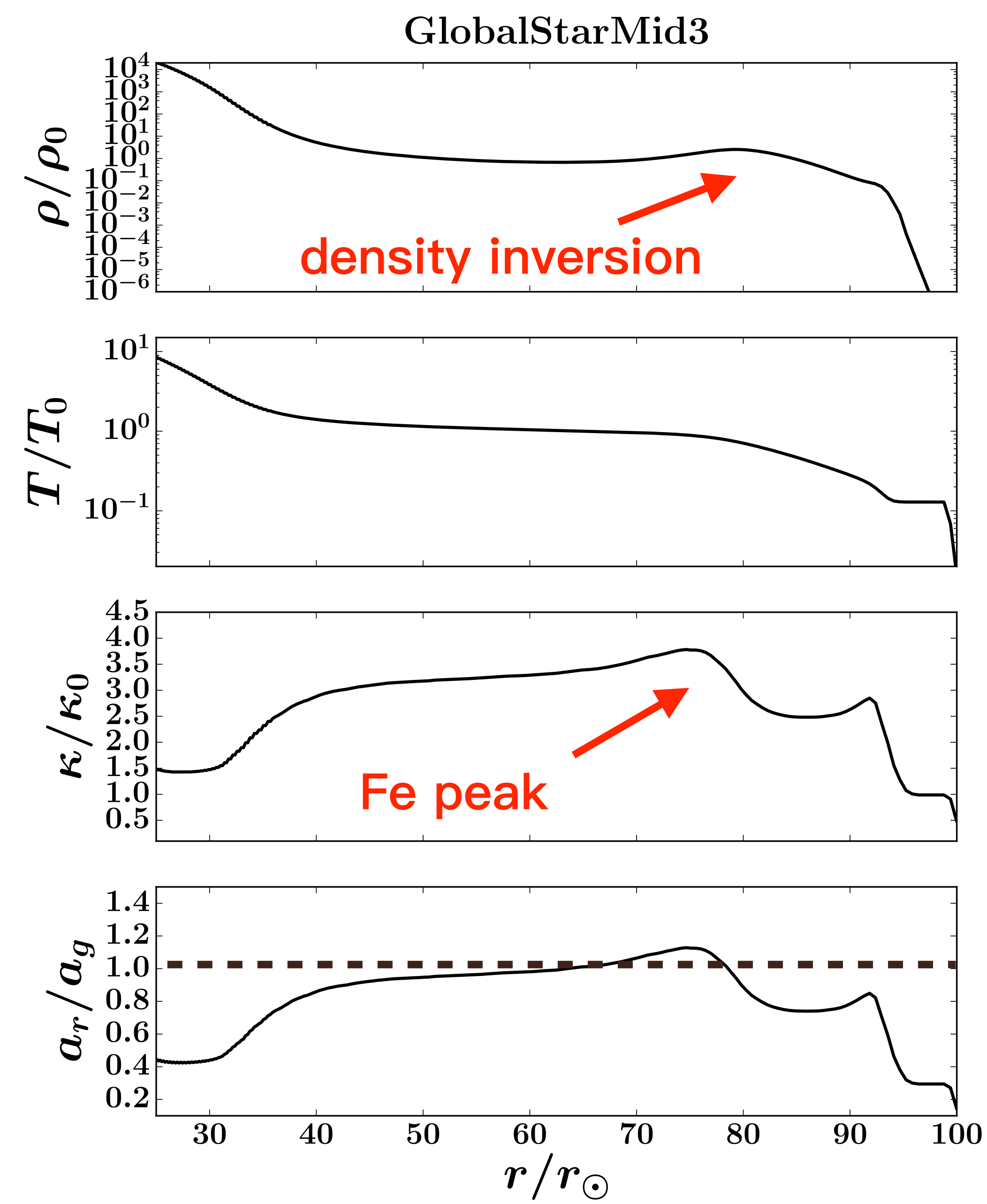

**PRELIMINARY RESULTS!**

**Jiang et al., in prep**

# Global Simulations of the Massive Star Envelopes

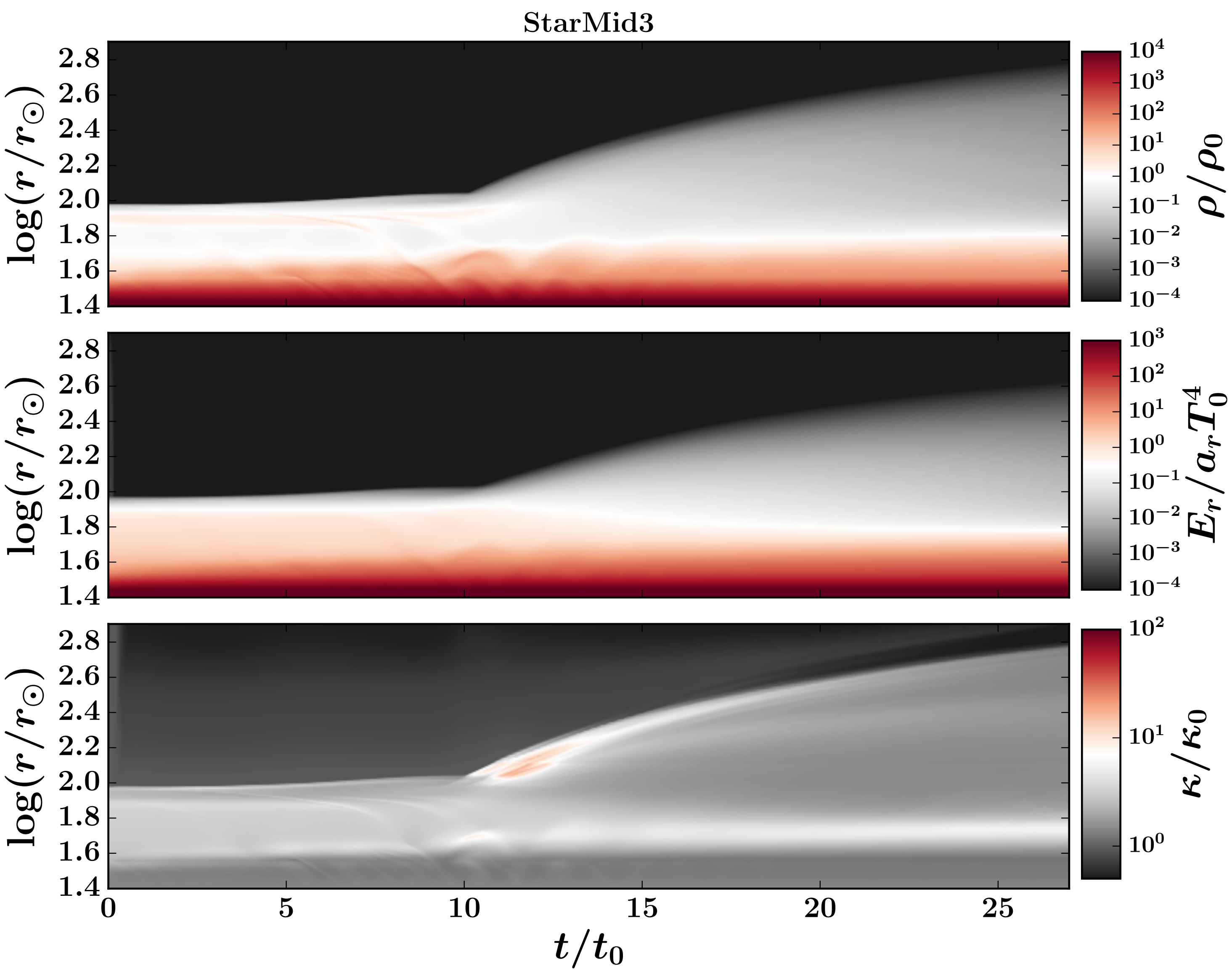



**The initial profile: Hydrostatic and thermal equilibrium**

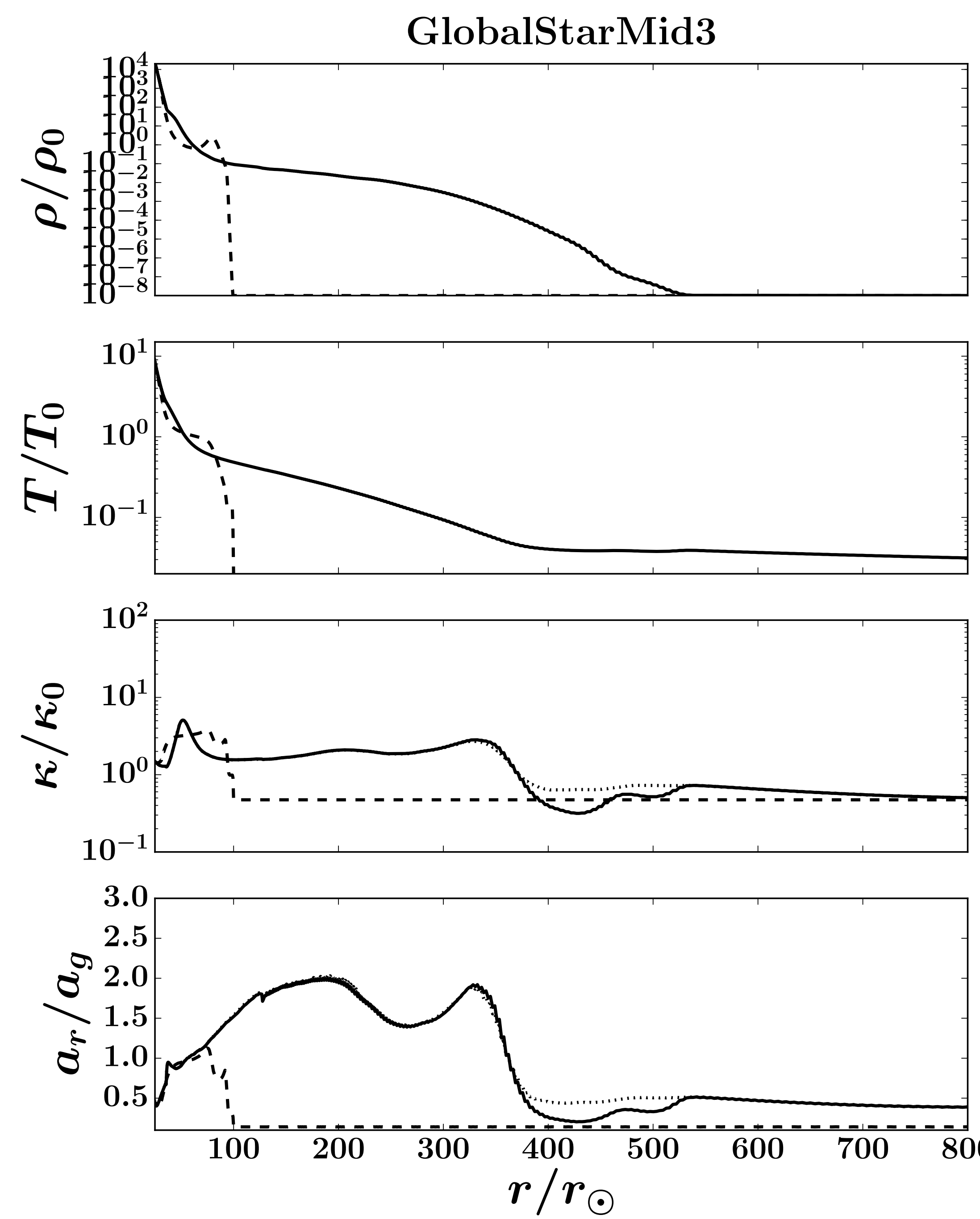



**PRELIMINARY RESULTS!**

# Global Simulations of the Massive Star Envelopes




Snapshot at $t = 18t_0$




$t_0 = H_0/c_{s,0}$

# Summary

- When $\tau_0 \gg \tau_c$ , convection is efficient and the simulations calibrate the mixing length theory with $\alpha = 0.55$
- When $\tau_0 \ll \tau_c$ , convection is inefficient and convection flux is much smaller than the predicted values by mixing length theory.
- The porosity factor reduces the effective radiation acceleration in the inefficient convection regime.
- Magnetic field reduces the stellar radius, increases the density fluctuation and the porosity factor.
- Preliminary results show the development of winds driven by the continuum radiation around the iron opacity peak.